

The Multiwavelength Properties of a Sample of Magellanic Cloud and Galactic Supersoft X-ray Binaries

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Abstract

Supersoft X-ray sources constitute a class of astronomical objects characterized by extremely high X-ray luminosities ($\sim 10^{37}$ erg s $^{-1}$), and low effective temperatures, typically 20-100 eV. The canonical model for these sources involves a white dwarf (WD) accreting from a more massive companion in a binary system at a very high rate ($\sim 10^{-7}$ M $_{\odot}$ yr $^{-1}$). The high soft X-ray luminosity is derived from steady or quasi-steady nuclear burning of accreted hydrogen in a shell on the WD surface. In this study, an observational investigation of accretion-related variability on time-scales of seconds to years is carried out for a selection of supersoft X-ray binaries. The principal target of interest is CAL 83 in the Large Magellanic Cloud (LMC). The other included targets are the LMC source RX J0513.9-6951 (RXJ0153), and MR Vel in the Milky Way. According to the literature, these sources have several properties in common. Each source contains a massive WD accreting through an accretion disc. Signatures of outflows with velocities of several thousand km s $^{-1}$ are present in their spectral line profiles in the form of Doppler shifted emission and P Cyg absorption features. They are non-eclipsing binaries, and the orbital inclinations of especially CAL 83 and RXJ0513 are very low. The latter two sources exhibit large-scale anti-correlated X-ray and optical variability on superorbital time-scales >100 d, which has been explained by cyclic changes of the WD envelope between a contracted high-temperature and an expanded low-temperature state. A variable ~ 67 s X-ray periodicity is found in the *XMM-Newton* lightcurves of CAL 83. This periodicity can be explained by a model similar to the LIMA model developed for dwarf nova oscillations. A correlation between X-ray temperature and flux is also confirmed on all the observed time-scales. New SALT spectra of the three sources are presented, and confirm the presence of outflows, which may be driven by magnetohydrodynamic (MHD) processes. Variable emission from ionized O VI appears to support the presumed temperature modulations associated with superorbital cycles. SHOC fast photometric lightcurves of these sources reveal no strictly periodic modulations, but rather quasi-periodic modulations on various time-scales of the order of 1000 s. As in cataclysmic variables, these phenomena are expected to be related to MHD turbulence. The OGLE-IV lightcurves of CAL 83 and RXJ0513 spanning the last 6 years confirm the continuation of superorbital cycles of the same nature as previously during the MACHO and OGLE-III projects.

Keywords: X-rays: binaries – binaries: close – white dwarfs – stars: oscillations – accretion, accretion discs – MHD – stars: winds, outflows – stars: individual: CAL 83 – stars: individual: RX J0513.9-6951 – stars: individual: MR Vel

Opsomming

Sagte X-straal bronne vorm 'n onderafdeling van astronomiese sisteme wat gekenmerk word deur geweldige hoë X-straal helderhede ($\sim 10^{37}$ erg s $^{-1}$), en lae temperatuur, tipies 20-100 eV. Die aanvaarde model vir hierdie bronne behels 'n wit dwerg in 'n binêre sisteem wat akkresie ondergaan teen 'n baie hoë tempo ($\sim 10^{-7}$ M $_{\odot}$ yr $^{-1}$) vanaf 'n geselster met 'n groter massa as die wit dwerg. Die oorsprong van die hoë sagte X-straal helderheid is die gelykmatige of kwasi-gelykmatige kernbranding van waterstof in 'n skil op die wit dwerg se oppervlak. In hierdie studie word 'n waarnemingsgeöriënteerde ondersoek uitgevoer na akkresieverwante veranderlikheid oor tydskele van sekondes tot jare in 'n versameling van sagte X-straal binêre sisteme. Die hoofbron in hierdie studie is CAL 83 in the Groot Magellaanse Wolk. Die ander bronne wat ingesluit is, is die bron RX J0513.9-6951 (RXJ0153), ook in die Groot Magellaanse Wolk, en MR Vel in die Melkweg. Volgens die literatuur, het hierdie bronne verskeie eienskappe in gemeen. Elke bron bevat 'n hoë-massa wit dwerg wat akkresie ondergaan deur 'n akkresieskyf. Bewyse van uitvloei teen snelhede van 'n hele paar duisend km s $^{-1}$ is teenwoordig in die profiele van die spektraallyne van hierdie bronne, in die vorm van Doppler-verskuifde emissie, sowel as P Cyg absorpsiekomponente. Geen eklips van een van die binêre komponente word waargeneem in hierdie bronne nie, en die orbitale inklinasiehoek van veral CAL 83 en RXJ0513 is baie klein. Laasgenoemde twee bronne openbaar grootskaalse X-straal en optiese variasie oor lang tydskele van >100 d, met 'n anti-korrelasie tussen die X-straal en optiese vloed. Hierdie verskynsel word verklaar deur sikliese veranderings van die wit dwerg atmosfeer tussen 'n gekrimpte hoë-temperatuur toestand, en 'n uitgesette lae-temperatuur toestand. 'n Veranderlike ~ 67 s ossillasie word gevind in die *XMM-Newton* ligkurwes van CAL 83. Hierdie ossillasie word verklaar deur 'n model soortgelyk aan die LIMA model wat ontwikkel is vir dwergnova ossillasies. 'n Korrelasie tussen X-straal temperatuur en -vloed word bevestig oor alle waargenome tydskele. Nuwe SALT spektra van die drie bronne word aangebied, en bevestig die teenwoordigheid van uitvloei, wat moontlik gedryf word deur magnetohidrodinamiese (MHD) prosesse. Dit kom voor asof veranderlike emissie van geïoniseerde O VI die vermoedelike temperatuurveranderinge gedurende die lang-tydskaal siklusse ondersteun. SHOC vinnige fotometriese ligkurwes van hierdie bronne openbaar geen streng periodiese ossillasies nie, maar eerder kwasi-periodiese modulاسies oor verskeie tydskele van die orde van 1000 s. Soos wat die geval is in kataklismiese veranderlikes, word daar verwag dat hierdie verskynsel geassosieer word met MHD turbulensie. Die OGLE-IV ligkurwes van CAL 83 en RXJ0513 wat oor die afgelope 6 jaar strek, bevestig voortsetting van langtermyn siklusse van dieselfde aard as wat voorheen waargeneem is tydens die MACHO en OGLE-III projekte.

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Chapter 1

Introduction

Population studies indicate that at least half of the stellar objects in the Universe are members of binary systems (e.g. Heintz, 1969; Herczeg, 1982). If the initial masses of the binary components are different, their evolution will occur on different time-scales, and one of the stars will evolve into a compact object, i.e. a white dwarf, neutron star or black hole, before its companion. During certain evolutionary stages of compact binaries, mass transfer takes place from the companion star to the compact object. This interaction results in the occurrence of a rich plethora of energetic phenomena, providing a cosmic laboratory for studying fundamental physical processes related to e.g. gravitation, viscosity and magnetic fields, as well as the associated radiation fields, under extreme conditions.

The sources investigated in this study belong to the class of supersoft X-ray binaries, which forms a distinct subclass of compact binaries. Therefore, a brief discussion of the various classes of compact binaries will be presented below, before shifting the focus to the specific properties of supersoft X-ray binaries.

The compact binary class consists of several different types of systems, most of which can be subdivided into three broad categories (e.g. the review of Van den Heuvel, 2009):

(i) A system in which one of the stars is a neutron star or a black hole, is commonly referred to as an *X-ray binary*. Mass transfer occurs from the companion, also called the “donor”, to the compact object. X-ray binaries are subdivided into low-mass X-ray binaries (LMXBs) in which the donor mass is $\leq 1 M_{\odot}$, and high-mass X-ray binaries (HMXBs) which have massive ($> 10 M_{\odot}$) donors, although systems with intermediate-mass donors are also known¹. An artist’s representation of an LMXB is provided in Fig. 1.1, but the general configuration in this figure can be applied to other accreting compact binaries as well.

¹ $M_{\odot}$ represents the solar mass.

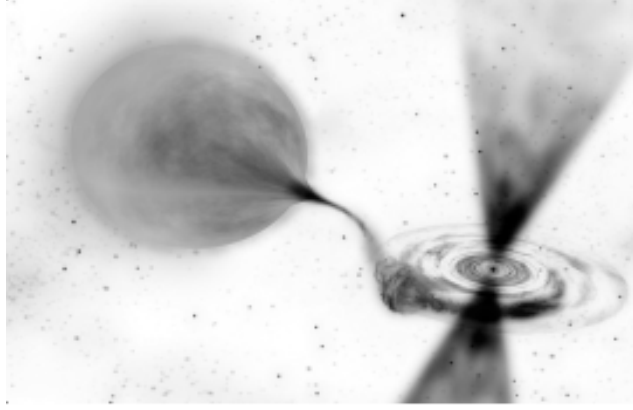


Fig. 1.1: An artist's representation of a low-mass X-ray binary. The neutron star on the right is undergoing accretion from the companion on the left through an accretion disc. The cone-like structures represent jets of material ejected from the inner disc regions. (Adopted from Seward and Charles, 2010, Fig. 11.39, p. 202, with image credit to Dana Berry, SkyWorks Digital/NASA-GSFC.)

(ii) The second class is the *double-degenerates*. A double-degenerate binary consists of two degenerate objects (white dwarf or neutron star) in a single system.

(iii) Compact objects containing a white dwarf with a non-degenerate companion include a wide variety of systems which can be referred to as *cataclysmic variable*-like binaries. The properties of the latter objects are of particular interest to this study.

A detailed treatment of cataclysmic variables can be found in Warner (1995*a*), while Hellier (2001) provides a more general introduction. As the name indicates, cataclysmic variables (CVs) are characterized by substantial brightness variations. The most dramatic variability is observed during nova outbursts, which constitute runaway nuclear burning of hydrogen-rich material accreted onto the surface of the white dwarf (WD). The brightness increase from the pre-nova state to maximum brightness, ranges from ~ 6 magnitudes up to ~ 19 magnitudes. CVs in which only one nova eruption has been observed, are called *classical novae*; if more than one outburst has been observed, the system is a *recurrent nova*. It will be shown in §2.6.3 that a nova outburst can occur when the mass of the accreted envelope on the WD surface reaches a certain critical value associated with temperatures and pressures high enough to trigger thermonuclear burning. During the outburst, a fraction of the accreted material is blown off the WD. During the outburst, a fraction of the accreted material is blown off the WD.

Some other CVs exhibit outbursts with much smaller amplitudes, typically 2 to 5 magnitudes. These outbursts are not caused by thermonuclear burning, but are the

result of a thermal instability in the accretion disc surrounding the WD, as explained in §2.7.4. CVs that exhibit these outbursts are referred to as *dwarf novae*.

CVs also include a group that is known as *nova-like variables* or *novalikes*, which are high accretion rate CVs (see below), in which no large eruptions as in novae or dwarf novae are observed. This may imply that the source is intrinsically non-eruptive, or simply that no eruptions have yet been *observed* in the source. This class therefore includes CVs in pre- or post-novae states. The typical mass transfer rates encountered in CVs range from $\sim(2-50)\times 10^{-11} \text{ M}_{\odot} \text{ yr}^{-1}$ in dwarf novae during quiescence, to $\sim(3-10)\times 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$ in nova remnants, novalikes and dwarf novae during outburst (e.g. Warner, 1995*a*, pp. 64-66). The typical luminosity of a nova remnant or novalike is $\sim 10^{33} \text{ erg s}^{-1}$, and is derived from the release of gravitational potential energy of material transferred from the companion undergoing accretion onto the WD.

During the last few decades, another class of WD binaries closely related to CVs has been established, i.e. *supersoft X-ray sources*. The first of these supersoft X-ray sources (SSSs) were discovered in the Large Magellanic Cloud (LMC) and the Small Magellanic Cloud (SMC) with the *Einstein Observatory* (Long, Helfand and Grabelsky, 1981; Seward and Mitchell, 1981). Further observations and new discoveries during the *ROSAT* all-sky survey (Trümper et al., 1991) and subsequent pointed observations established these sources as a distinct new class of objects. The defining observational properties of this class are (i) their extreme X-ray luminosities, ranging from $\sim 10^{36} \text{ erg s}^{-1}$ to as high as the Eddington limit of $\sim 10^{38} \text{ erg s}^{-1}$, and (ii) the extreme softness of their X-ray spectra, typically with effective temperatures of $kT_{\text{eff}} \sim 20\text{-}100 \text{ eV}$.

The vast majority of observed SSSs are binary systems. Van den Heuvel et al. (1992) showed that the X-ray properties can be explained by steady nuclear burning of accreted hydrogen on the surface of a WD accreting via Roche lobe overflow. To sustain persistent hydrogen nuclear burning on the WD surface, the mass transfer rate in SSSs needs to be significantly higher than in CVs, and has been shown to be of the order of $10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}$ (see §2.6.3). In CVs, the WD is more massive than the donor. However, the high accretion rate required to drive persistent surface nuclear burning in SSSs is only possible if the donor is more massive than the WD. The main characteristics of SSSs distinguishing them from CVs are therefore their inverted mass ratios, high accretion rates and extreme luminosities derived from surface nuclear burning.

The model of Van den Heuvel et al. (1992) is often referred to as the *close binary supersoft source* or CBSS model, and is explained further in Chapter 2. The three

targets studied in this thesis form part of the CBSS class. However, SSSs do not form a homogeneous class of objects, and other variants are discussed in §2.9. A few years ago, the census of SSS were: ~ 15 in the Milky Way, 18 in the LMC (Kahabka et al., 2008), 6 in the SMC and 120-160 in about 20 external galaxies (Kahabka, 2006), but several new discoveries have since also been reported.

Soft X-rays are subject to significant interstellar extinction, as explained in §2.6.5. Thus, the number of SSSs discovered in the Milky Way is quite low due to the high amount of absorption in the Galactic plane. Sources in other galaxies can also only be observed if they are situated on the near side of that galaxy, near the outer edge of the interstellar hydrogen layer (e.g. Kahabka and van den Heuvel, 2006). For example, the actual population of SSSs in M31 and the Milky Way is estimated to be approximately 2 orders of magnitude more than the number that is actually observed (Rappaport, Di Stefano and Smith, 1994).

Although the SSSs initially described by the Van den Heuvel et al. (1992) model were expected to be reasonably steady sources, it has since been found that several of the sources exhibit significant variability on various time-scales. Previous authors have shown that the supersoft X-ray sources CAL 83 and RX J0513.9-6951 (hereafter RXJ0513) in the LMC exhibit cyclic transitions between optical low and optical high states on superorbital time-scales of >100 d, as well as X-ray variability that is anti-correlated with the optical variability. This phenomenon has been explained by limit cycle models involving sustained transitions between periods of enhanced and reduced mass transfer.

The question of the nature of the progenitors of Type Ia supernovae (SNIa) has been a controversial one in astronomy circles for quite some time. Although many observational and theoretical studies of CVs in the past seemed to indicate that virtually all accreted material is blown off the WD surface during a nova outburst, recent results of e.g. Starrfield (2015) indicate that the WD in a CV can indeed be expected to grow in mass.

In any case, even though the nuclear burning on the WD surface may be variable in SSSs, these sources do not show the violent outbursts observed in novae. This makes it even more likely that most of the accreted mass (including the nuclear burning “ashes”) may remain on the WD surface, increasing the WD mass over time. The recognized high rate of mass accretion in SSSs without sizeable mass loss, implies that at least some SNIa may originate from SSSs (e.g. Podsiadlowski, 2010). A better understanding of accretion and outflows in SSSs can therefore play an important role in solving this fundamental problem.

Evidence for the existence of high-velocity outflows with velocities of the same order as the escape velocity of a WD has been found in 4 supersoft X-ray binaries. The galactic sources RX J0019.8+2156 and RX J0925.7-4758 (also known as MR Vel), as well as RXJ0513, exhibit sharp blue- and redshifted “satellite” features mirrored on the main Balmer emission lines (and He II in some cases) (Southwell et al., 1996a; Becker et al., 1998; Tomov et al., 1998; Motch, 1998). CAL 83 exhibits broad, variable Doppler shifted wing structures that should also be associated with an outflow, albeit a “wider” one with a larger collimation angle (e.g. Crampton et al., 1987). P Cygni profiles in the optical H I lines, as well as in some X-ray lines, serve as additional evidence of mass loss in these systems.

Observational evidence indicates that the WD masses in these 4 systems are approximately $1.3 M_{\odot}$, just below the Chandrasekhar limit of $1.4 M_{\odot}$ (see Chapter 3, as well as Kahabka, 1995). This may indicate that these sources have already accreted a substantial amount of mass from the secondary, and may thus have been in the supersoft soft phase for a substantial amount of time. The UV and optical spectra of these systems indicate the presence of an irradiated accretion disc around the WD. A natural consequence of an extended period of high-rate disc accretion is that the WD may be spun up by disc torques to a short rotation period. However, no WD rotation periods have been published previously for these sources.

The question now arises as to the cause of these outflows. As will be discussed in §2.8, these outflows may be driven by continuum radiation, but magnetically driven winds or collimated outflows are also expected to occur in systems where magnetohydrodynamic (MHD) turbulence plays a significant role. It is well known that molecular viscosity alone is several orders of magnitude too small to drive the transport of angular momentum through an accretion disc (e.g. Frank et al. 2002, pp. 69-70), and especially in supersoft X-ray binaries with their high accretion rates, MHD turbulent viscosity is expected to play a significant role.

In addition to the strictly periodic WD rotation periods observed in some CVs, these systems are also known to exhibit erratic flickering, as well as quasi-coherent modulations, known as dwarf nova oscillations (DNOs) and quasi-periodic oscillations (QPOs) (§2.7.4). These oscillations are believed to be related to MHD processes occurring in the accretion disc. However, these phenomena have not been studied well in SSSs. We expect similar periodicities to be found in the lightcurves of SSSs, although the substantially higher accretion rate may result in somewhat different periodicity characteristics.

The aim of this study is to investigate the nature of the accretion process in a selection of supersoft X-ray binaries by studying the multiwavelength (X-ray to

optical) variability in these sources on time-scales of ~ 1 second to several years, with special consideration of the relevance of MHD turbulence. The main target of this study is the prototypical supersoft X-ray binary CAL 83, although new data of the other similar southern sources RXJ0513 and MR Vel are also presented. RX J0019.8+2156 was initially not selected for the current study due to its northern declination and associated inaccessibility to SALT, but will be included in future studies. In the light of the discussions above, the focus areas of this study within the main theme are the following:

- Searching for short time-scale periodicities from the WD break-up period, and higher in the X-ray and optical lightcurves that may be associated with the rotation period of a spun-up WD, as well as investigating the stability of any such periodicities, and evaluating the implications related to the intensity of the WD magnetic field.
- Searching for signatures of (possibly magnetocentrifugal) outflows in the optical spectra, and utilizing these to evaluate the properties of the outflows in these systems.
- Searching for quasi-coherent modulations in the optical lightcurves indicative of processes related to MHD turbulence in the accretion disc.
- Utilizing the long-term photometric lightcurves of CAL 83 and RXJ0513 to provide an update on the superorbital modulations.

The data and observational techniques employed to achieve these goals are the following:

- Timing analysis of archival *XMM-Newton* X-ray lightcurves.
- Spectral analysis of new optical spectra obtained with the Southern African Large Telescope (SALT).
- Timing analysis of photometric data from the Sutherland High-Speed Optical Cameras (SHOC), the MAssive Compact Halo Objects (MACHO) Project, and the Optical Gravitational Lensing Experiment (OGLE).

This thesis is structured as follows: In Chapter 2, an overview of the relevant astrophysical processes occurring in supersoft X-ray binaries is provided. This is followed by a detailed literature review of CAL 83, RXJ0513 and MR Vel in Chapter 3. This literature review is intended to provide a comprehensive background regarding the properties of these sources for the current study, but also to serve as a reference to aid in planning the follow-up work on these systems. The results obtained from *XMM-Newton* data of CAL 83 are presented in Chapter 4, followed by results of the spectral analysis of new SALT spectra of CAL 83, RXJ0513 and MR Vel in Chapter 5. In Chapter 6, SHOC lightcurves of the three sources will be presented, as well

as the newest OGLE lightcurves of CAL 83 and RXJ0513. A list of physical constants is provided in Appendix A. The tables containing the *XMM-Newton* Optical Monitor magnitudes discussed in Chapter 4 are presented in Appendix B. The general properties of spectral lines, which are applicable to Chapter 5, are summarized in Appendix C. A summary of the properties of CCD cameras and the processes applicable to reducing CCD data is given in Appendix D, followed by a discussion of the differential photometry technique that was used to correct the SHOC lightcurves (Appendix E). Finally, the publication list of the author is provided in Appendix F, with the first-author papers attached.

Chapter 2

Supersoft X-ray binaries

In this chapter, an overview of the astrophysical processes in supersoft X-ray binaries is presented. The purpose of this chapter is not by any means to provide a comprehensive discussion or review of the wide range of topics covered here, but rather to highlight the relevant theoretical concepts and their observational significance to provide a framework to evaluate the research presented in this thesis.

Although much of the theory summarized here is applicable to all accreting compact binaries, the focus here will be on accreting white dwarfs in supersoft X-ray binaries, but also in the very closely related class of cataclysmic variables (CVs). In terms of notation, the convention will be to use the symbols M_{wd} and R_{wd} for the mass and radius of the white dwarf if the relevant equation is specifically applicable to a white dwarf, and the symbols M_1 and R_1 for equations that are generally applicable if the primary is a compact object like a neutron star, or, in some cases, a black hole.

The chapter starts off with a precursory look at the by-now canonical model for close binary supersoft X-ray sources (CBSSs) of Van den Heuvel et al. (1992), followed by several sections in which certain aspects of the processes taking place in these systems are explored in more detail: In §2.2, a basic introduction to binary stars is provided, followed by three sections elaborating on the process of accretion onto a white dwarf, and a section on the properties of the accreting white dwarf itself. In §2.7 and §2.8, two related topics that are especially relevant for this work are discussed: the occurrence of turbulence in, and outflows from, accretion discs. The chapter concludes with some remarks on the different types of supersoft X-ray sources that are distinguishable, as well as their importance as a phase in the bigger picture of compact binary evolution.

2.1 The model for close binary supersoft X-ray sources

In the early years, it was suggested that supersoft X-ray sources (SSSs) might be low-mass X-ray binaries (LMXBs), i.e. binaries containing an accreting neutron star (Greiner, Hasinger and Kahabka, 1991) or a black hole (Smale et al., 1988; Cowley et al., 1990) (see the related discussions in the literature review on CAL 83 in §3.1). Following the initial tentative interpretations, numerical calculations of the resulting spectra when a neutron star accretes at a rate close to the Eddington limit were performed, and it was found that this scenario may indeed be associated with the emission of supersoft X-rays (Kylafis and Xilouris, 1993; Kylafis, 1996).

However, Van den Heuvel et al. (1992) proposed an alternative model in which a white dwarf (WD) accretes material from a possibly slightly evolved binary companion at a very high rate, with the nuclear burning of accreted material on the WD surface producing the high soft X-ray luminosity. Although the nature of the compact object is not quite confirmed in many SSSs, observational evidence makes the Van den Heuvel et al. (1992) model the favoured one for many of these sources.

Van den Heuvel et al. (1992) provided some compelling arguments against the interpretation of SSSs as LMXBs. These authors pointed out that the effective temperatures of the known black hole candidates are still of the order of a few keV, while SSS temperatures are in the range of a few tens of electronvolts. While Greiner et al. (1991) argued that an increasing mass transfer rate onto a neutron star will cause a cocoon of material to build up around the neutron star, causing an increased photospheric radius and softer X-ray spectrum, observational evidence does not seem to support this.

In the neutron star binary Cen X-3, accreting near the Eddington limit, the X-ray spectrum becomes harder when the accretion rate increases, and the X-rays are quenched for super-Eddington accretion rates (Giacconi, 1975). The quenching occurs due to the photoionization of surrounding matter by the X-rays from the central source, “degrading” the X-rays to optical and UV emission. However, in the source LMC X-4, the luminosity does go up to ~ 20 times the Eddington value, but the effective temperature remains a few keV (Levine et al., 1991).

Instead, Van den Heuvel et al. (1992) proposed that the accreting object may be a WD with mass $0.7\text{--}1.3 M_{\odot}$. The photospheric radii derived from blackbody fits to SSS X-ray spectra are in very good agreement with what one would expect for a hydrogen burning layer on the surface of a WD. The X-ray emission observed from the nova GQ Mus (Ögelman et al., 1993) and the symbiotic star SMC 3 (Kahabka, Pietsch and Hasinger, 1994) provided important observational evidence that super-

soft X-ray emission can indeed be associated with thermonuclear burning on the surface of a WD.

In the accreting WD model, the accretion disc is the dominant source of optical light, and is much brighter than the unheated side of the donor. A stellar wind is expected to blow from the heated side of the donor, and Van den Heuvel et al. (1992) argued that the He II $\lambda 4686$ emission line originates from the interaction between this wind and a wind from the WD/disc, and that its radial velocity semi-amplitude is much smaller than the actual value for the WD. If this is the case, it will lead to a severe underestimation of the mass function, and therefore an overestimation of the primary mass, pointing, erroneously in most cases, towards neutron star or black hole primaries (see e.g. the discussion in §3.1).

It will be shown in §2.6.3 that accretion rates of the order of $\sim 10^{-7} M_{\odot} \text{ yr}^{-1}$ can drive steady nuclear burning on the surface of a WD. Such a high accretion rate can be sustained by Roche-lobe overflow in a short-period WD binary if the donor's Roche lobe (and by implication also the binary separation), is shrinking, yielding thermally unstable mass transfer. In the conservative case, this requires that the secondary mass be similar to or greater than the primary mass (Paczynski, 1971; Savonije, 1983, see also Eq. (2.19)). The computational results of Pylyser and Savonije (1988, 1989) showed that accretion rates in the steady burning range is obtained for companion masses in the 1.5-2 $M_{\odot}$ range, with the accretion rate depending heavily on the mass and evolutionary state of the donor, and only slightly on the WD mass (for M_{wd} in the 0.7-1.3 $M_{\odot}$ range).

Since the mass transfer and resultant accretion rate in compact binaries are intimately connected to the binary orbital parameters and binary evolution, a brief discussion of these will be presented, focussing only on the most relevant aspects related to this study.

2.2 Binary parameters

In a binary system, two stars are continuously orbiting their common centre of mass. Their motion is governed by Kepler's third law, given by (e.g. Carroll and Ostlie, 1996, Chapter 7)

$$G(M_1 + M_2)P_{\text{orb}}^2 = 4\pi^2 a^3 . \quad (2.1)$$

Here G is the gravitational constant¹, M_1 and M_2 are the stellar masses and P_{orb} the orbital period. The semi-major axis of the orbit of the reduced mass is $a = a_1 + a_2$ (where a is also known as the binary separation), with a_1 and a_2 the semi-major axes

¹A list of physical constants is provided in Appendix A.

of the orbits of the two components about the centre of mass (CM) of the system.

Due to the Doppler effect, relative movement between a source of radiation and the observer will cause a shift in the observed wavelength of the radiation. If the source is moving towards the observer, a shorter or “blueshifted” wavelength is observed, while if the source is moving away from the observer, a longer or “redshifted” wavelength is observed. For non-relativistic velocities, this only applies to movement directed along the line of sight, i.e. *radial* movement. The observed wavelength λ_{obs} is related to the radial velocity v and the rest wavelength λ_0 by

$$\frac{\lambda_{\text{obs}} - \lambda_0}{\lambda_0} = \frac{v}{c}, \quad (2.2)$$

where c is the speed of light in vacuum. Different manifestations of the Doppler effect are very powerful diagnostic tools in astronomical spectroscopy, especially in compact binary systems. Specifically, when observing a spectral line from a star in a binary system, the wavelength of the line will continuously oscillate about the rest wavelength as the star moves along its orbital path, and the period of the oscillation will be the orbital period. If it is assumed that the orbits of the stars are circular (which is often a very accurate assumption), then the speeds of the two stars are constant and are given by

$$v_1 = \frac{2\pi a_1}{P_{\text{orb}}} \quad \text{and} \quad v_2 = \frac{2\pi a_2}{P_{\text{orb}}} \quad (2.3)$$

respectively.

Only the radial component of the relative movement will result in a Doppler shift, bringing us to another important parameter to consider when dealing with binary systems: the inclination angle i of the system. The inclination is defined as the angle between the normal to the orbital plane and the line of sight. Therefore, an eclipsing system will have an inclination angle of $\sim 90^\circ$, while $i = 0^\circ$ when the orbital plane is viewed directly face-on. The maximum radial velocity for each star respectively during its orbital motion therefore differs from the absolute velocities given in Eq. (2.3) by a factor of $\sin i$:

$$K_1 = v_1 \sin i \quad \text{and} \quad K_2 = v_2 \sin i. \quad (2.4)$$

K_1 and K_2 are known as the radial velocity semi-amplitudes of the two stars respectively, as each of them represents the semi-amplitude of the sinusoidal curve that will be obtained when plotting radial velocity measurements versus time for the star. It is evident that, if $i = 0^\circ$, the radial velocity semi-amplitude will be zero.

Independent of the orbital motion, there is also the radial velocity of the system as a whole (i.e. the centre of mass of the system) relative to the observer. This is known as the *systemic velocity*.

Utilizing the definition of the centre of mass, as well as Eq. (2.3) and Eq. (2.4), Kepler's third law in Eq. (2.1) can be rewritten as

$$\frac{M_2^3}{(M_1 + M_2)^2} \sin^3 i = \frac{P_{\text{orb}}}{2\pi G} K_1^3 . \quad (2.5)$$

The right hand side of this equation is known as the *mass function*, and can obviously be obtained if the orbital period of the system and the radial velocity semi-amplitude of one of the components are known. If, for example, the component masses and mass function is known, one can constrain the inclination. Conversely, if the inclination is known, the mass function allows one to determine the ratio $q = M_2/M_1$ of the component masses. However, for a known inclination, absolute values for the component masses can be obtained only if one of the masses was determined with an independent method (e.g. by determining its spectral class), or if the radial velocity semi-amplitude of the other star (K_2) is also known.

2.3 Mass transfer and accretion

In compact binary systems, it is often observed that mass is transferred from one of the stars, called the *secondary*, to the other star, called the *primary*, and is accreted onto the surface of the primary. The basic mechanism of this process is summarized in this section, which is based on the approach of Frank, King and Raine (2002, pp. 1-7, 48-58), with reference to other sources where indicated.

2.3.1 The Roche lobes

Any test particle in the vicinity of a binary system will be subject to the combined gravitational attraction from the two stars. In order to understand the resultant gravitational field acting on such a test particle, it is helpful to make use of the Roche approach. In this approach, the two stars are assumed to act like point masses executing circular Keplerian orbits around each other in the orbital plane. The stars are so massive that the test particle will not perturb their combined gravitational field. Within a reference frame rotating together with the binary system, at angular velocity ω relative to an inertial frame, the effect of gravitational and centrifugal forces are combined in an effective Roche potential, given by

$$\Phi_{\text{R}}(\mathbf{r}) = -\frac{GM_1}{|\mathbf{r} - \mathbf{r}_1|} - \frac{GM_2}{|\mathbf{r} - \mathbf{r}_2|} - \frac{1}{2}(\omega \times \mathbf{r})^2 . \quad (2.6)$$

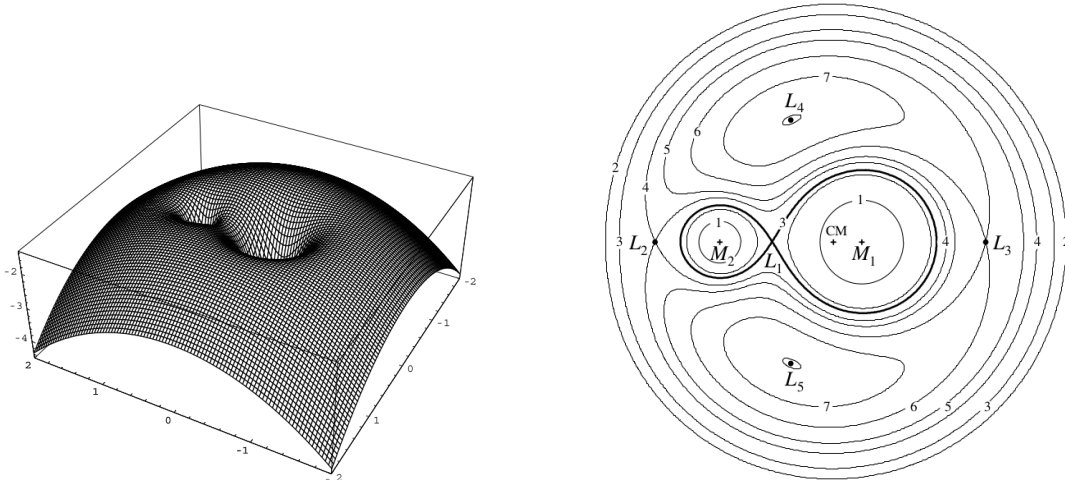


Fig. 2.1: Left: Roche potential surface for $q = 0.25$. Right: Its sections in the orbital plane for Φ_R constant. The labels 1 to 7 indicate surfaces of increasing Φ_R . Also shown are CM (the centre of mass) and the Lagrange points $L_1 - L_5$. (Adopted from Frank et al., 2002, Fig. 4.2 and 4.3, pp. 51-52.)

Here $\mathbf{r}$ represents the position vector of the test mass from the CM of the system, while the vectors $\mathbf{r}_1$ and $\mathbf{r}_2$ extend from the CM of the system to the centres of the two stars respectively.

In Fig. 2.1, the Roche potential for a binary system with $q = 0.25$ is shown, as well as a cross-section in the orbital plane, with the contour lines representing equipotential surfaces. The figure-of-eight curve is especially important, and forms a teardrop-shaped lobe around each star in three dimensions, known as the star's *Roche lobe*. Within the star's Roche lobe, material is primarily under the influence of this particular star and can be considered to be gravitationally bound to the star. The point L_1 connecting the two Roche lobes is known as the inner Lagrange point and is a saddle point of the Roche potential. Material inside one of the Roche lobes, close to L_1 , will find it easier to move through L_1 into the other lobe than to escape the critical surface entirely.

Depending on the extent to which the two components fill their respective Roche lobes, binary systems can be subdivided into the following three groups:

- *Detached binary:* Both stars are smaller than their Roche lobes.
- *Semi-detached binary:* Only one of the stars is filling its Roche lobe, either because the star is expanding as part of its evolution, or because its Roche lobe is shrinking. Thermal motion can then push gas across the L_1 point to the Roche lobe of the primary, and this is known as Roche lobe overflow.
- *Contact binary:* Both stars are filling or over-filling their Roche lobes.

The radius $R_{L,2}$ of the Roche lobe of the secondary star is given by the approximate formula of Eggleton (1983),

$$\frac{R_{L,2}}{a} = \frac{0.49q^{2/3}}{0.6q^{2/3} + \ln(1 + q^{1/3})} , \quad (2.7)$$

which represents the radius of a sphere with volume equal to that of the teardrop-shaped lobe. An expression for the primary Roche lobe radius ($R_{L,1}$) can simply be obtained by replacing q with q^{-1} in the equation above. The fitted formula of Plavec and Kratochvil (1964) gives the distance b_1 from the centre of the primary to the L_1 point:

$$\frac{b_1}{a} = 0.500 - 0.227 \log q , \quad (2.8)$$

and can be used for $0.1 < q < 10$ (e.g. Hellier, 2001, p. 24).

In binary systems, the two main mechanisms through which mass transfer occurs are Roche lobe overflow and stellar winds from the secondary star. The latter is commonly known as Bondi-Hoyle-Lyttleton accretion (Hoyle and Lyttleton, 1939; Bondi and Hoyle, 1944), and involves the motion of the accreting primary through a wind from the donor, where the velocity of the wind is much larger than the primary orbital velocity (see also the summary in Edgar, 2004). Abate et al. (2013) also discussed a wind mass transfer mechanism that falls “between” Roche lobe overflow and Bondi-Hoyle-Lyttleton accretion, called wind Roche-lobe overflow (WRLOF). WRLOF involves very efficient accretion by the primary from a slower, dense wind from the secondary, as found in asymptotic giant branch stars. However, the current discussion will focus primarily on Roche lobe overflow, which is believed to be the mechanism responsible for mass transfer in most compact binaries.

During the process of mass transfer, the mass ratio q of the system will also change. From the preceding discussion, it is evident that a change in mass ratio will necessarily bring about a change in the orbital parameters and the Roche lobe configuration, and also on the mass transfer rate, which will in turn shape the further evolution of the system.

2.3.2 Roche lobe overflow and the role of angular momentum

Because of the orbital motion of the two components, mass transfer is associated not only with a redistribution of mass, but also with a redistribution of angular momentum. It is easy to show that the orbital angular momentum of the binary system is given by (Wynn and King, 1995, see e.g. Meintjes, 2002 for a review)

$$J_{\text{orb}} = M_1 M_2 \left(\frac{Ga}{M} \right)^{1/2} , \quad (2.9)$$

where the total mass of the binary system is $M = M_1 + M_2$. After logarithmically differentiating with respect to time, the rate of change of the binary separation is obtained:

$$\frac{\dot{a}}{a} = 2 \frac{\dot{J}_{\text{orb}}}{J_{\text{orb}}} + \frac{\dot{M}}{M} - 2 \frac{\dot{M}_1}{M_1} - 2 \frac{\dot{M}_2}{M_2} . \quad (2.10)$$

Because the secondary is losing mass, $\dot{M}_2 < 0$. Due to Roche lobe overflow, angular momentum is transferred from the secondary to the Roche lobe of the primary at a rate of

$$\dot{J}_{\text{ov}} = \dot{M}_2 (GM_1 R_{\text{circ}})^{1/2} , \quad (2.11)$$

where R_{circ} is the circularization radius, which is the distance from the primary where the ballistic stream has the same angular momentum as the L_1 point (explained in more detail in §2.4.1). However, not all of the mass and associated angular momentum that is lost by the secondary through the L_1 point will necessarily be accreted onto the primary. Material may also be ejected from the system, e.g. in a wind from the accretion disc or the ejection of accreting blobs of material from a fast rotating magnetic field associated with the primary star (see §2.8). Suppose that a fraction η of $\dot{J}_{\text{ov}}$ is lost from the binary, then the associated rate of angular momentum loss is

$$\frac{\dot{J}_{\text{loss}}}{J_{\text{orb}}} = \eta \frac{\dot{J}_{\text{ov}}}{J_{\text{orb}}} = \eta \dot{M}_2 \frac{(GM_1 R_{\text{circ}})^{1/2}}{J_{\text{orb}}} . \quad (2.12)$$

In addition to the angular momentum losses due to the ejection of material as part of the accretion process, angular momentum losses through other mechanisms are also possible, such as magnetic braking of the secondary (Mestel, 1968; Mestel and Spruit, 1987) or a stellar wind from the secondary, as well as gravitational radiation (Paczynski, 1967) or any other process through which the orbital angular momentum is drained. Denoting these other angular momentum losses by $\dot{J}_o$, the total rate of change of orbital angular momentum can be expressed as

$$\frac{\dot{J}_{\text{orb}}}{J_{\text{orb}}} = \frac{\dot{J}_{\text{loss}}}{J_{\text{orb}}} + \frac{\dot{J}_o}{J_{\text{orb}}} . \quad (2.13)$$

As

$$\dot{M} = \dot{M}_1 + \dot{M}_2 , \quad (2.14)$$

the mass loss rate of the secondary will be

$$-\dot{M}_2 = \dot{M}_1 - \dot{M} . \quad (2.15)$$

Suppose that a fraction α of the mass lost by the secondary through Roche lobe

overflow is lost from the system, then

$$\dot{M} = \alpha \dot{M}_2 \quad (2.16)$$

$$\dot{M}_1 = -(1 - \alpha) \dot{M}_2 . \quad (2.17)$$

By using Eq. (2.7), the rate of change of the secondary Roche lobe radius can now be expressed as

$$\begin{aligned} \frac{\dot{R}_{L,2}}{R_{L,2}} = \frac{-\dot{M}_2}{M_2} \left\{ 2 - \frac{1 + [1 - \alpha] q}{3} \left[2 - \frac{1.2q^{2/3} + 1 - \frac{1}{1+q^{1/3}}}{0.6q^{2/3} + \ln(1 + q^{1/3})} \right] \right. \\ \left. - 2[1 - \alpha] q - \frac{q\alpha}{1 + q} - 2\eta \frac{M_2 [GM_1 R_{\text{circ}}]^{1/2}}{J_{\text{orb}}} \right\} + \frac{2\dot{J}_o}{J_{\text{orb}}} . \end{aligned} \quad (2.18)$$

In the fully conservative case, with no loss of mass or angular momentum, the equation above reduces to

$$\frac{\dot{R}_{L,2}}{R_{L,2}} = \frac{-\dot{M}_2}{M_2} \left\{ 2 - 2q - \frac{1 + q}{3} \left[2 - \frac{1.2q^{2/3} + 1 - \frac{1}{1+q^{1/3}}}{0.6q^{2/3} + \ln(1 + q^{1/3})} \right] \right\} . \quad (2.19)$$

The factor $\frac{-\dot{M}_2}{M_2}$ has a positive value. For $q < 0.79$, the value of the term in curly brackets will be positive, and the secondary Roche lobe will expand. However, when $q > 0.79$, the term in the curly brackets is negative, and the secondary Roche lobe will shrink. The effect of the Roche lobe shrinking down on the donor will cause a very high mass transfer rate on the thermal time-scale of the donor, as is found in steady supersoft X-ray sources according to the Van den Heuvel et al. (1992) model.

At this stage it is worth taking note of the different time-scales influencing mass transfer, applicable to the secondary star, that are often referred to in binary evolution (e.g. Verbunt, 1993):

- The nuclear time-scale is given by

$$\tau_{\text{nuc}} \approx 10^{10} \left(\frac{M_2}{M_\odot} \right) \left(\frac{L_2}{L_\odot} \right)^{-1} \text{ yr} , \quad (2.20)$$

where $L_\odot$ is the solar luminosity. This is the time-scale on which the star expands due to hydrogen burning in its core.

- The thermal time-scale,

$$\tau_{\text{th}} \approx 3.1 \times 10^7 \left(\frac{M_2}{M_\odot} \right)^2 \left(\frac{R_2}{R_\odot} \right)^{-1} \left(\frac{L_2}{L_\odot} \right)^{-1} \text{ yr} , \quad (2.21)$$

represents the time-scale on which the star tries to restore perturbed thermal equilibrium.

- The time-scale on which a star restores its perturbed hydrostatic equilibrium is the dynamical time-scale, i.e.

$$\tau_{\text{dyn}} \approx 0.04 \left(\frac{M_2}{M_\odot} \right)^{-1/2} \left(\frac{R_2}{R_\odot} \right)^{3/2} \text{ d} . \quad (2.22)$$

Considering the typical Roche lobe geometry in a WD binary, $\sim 5\%$ of the radiation from the WD will strike the donor, which represents an amount of energy approximately two orders of magnitude more than what is typically emitted by the donor (e.g. Di Stefano and Nelson, 1996). This irradiation is likely to excite a strong wind from the donor. As an alternative to the high mass ratio q suggested in the CBSS model, Van Teeseling and King (1998) therefore suggested wind-driven evolution for SSSs containing low-mass secondary stars. A considerable loss of angular momentum in the wind from the secondary would then drive Roche lobe overflow at a rate high enough to sustain steady or recurrent hydrogen shell burning, even for $q \lesssim 0.7$. This mechanism was proposed for SMC 13, and also for the LMC SSS RX J0537.7-7034 (Greiner, Orio and Schwarz, 2000).

2.3.3 Accretion luminosity

The energy yield from accretion

Consider a mass m_{acc} accreting from afar (a distance $\gg R_1$) onto a star with mass M_1 and radius R_1 . The gravitational potential energy released will be

$$\Delta E_{\text{acc}} = \frac{GM_1 m_{\text{acc}}}{R_1} , \quad (2.23)$$

with a corresponding accretion luminosity of

$$L_{\text{acc}} = \frac{GM_1 \dot{m}_{\text{acc}}}{R_1} \\ \sim 5 \times 10^{36} \left(\frac{M_1}{M_\odot} \right) \left(\frac{\dot{m}_{\text{acc}}}{3 \times 10^{-7} M_\odot \text{ yr}^{-1}} \right) \left(\frac{R_1}{5 \times 10^8 \text{ cm}} \right)^{-1} \text{ erg s}^{-1} , \quad (2.24)$$

assuming that all the gravitational potential energy of the accreting material is converted to radiation at the star’s surface. It is evident that ΔE_{acc} is directly proportional to the “compactness” M_1/R_1 of the primary star. The radius R_{wd} of a zero-temperature WD is related to its mass M_{wd} by (Hamada and Salpeter, 1961;

Eracleous and Horne, 1996)²

$$R_{\text{wd}} = 4.9 \times 10^8 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right)^{-0.8} \text{ cm} . \quad (2.25)$$

according to which the WD radius decreases with increasing mass. For a WD with mass $\sim 1 M_{\odot}$ and radius $\sim 4.9 \times 10^8$ cm, the energy yield from accretion is $\sim 2.7 \times 10^{17}$ erg g⁻¹. This can be compared with the energy yield during the nuclear fusion to helium of a mass m consisting of hydrogen, which is

$$\Delta E_{\text{nuc}} = 0.007mc^2 . \quad (2.26)$$

This is equivalent to 6.3×10^{18} erg g⁻¹, which is more than an order of magnitude larger than the energy yield from accretion. This is why, in the case of supersoft X-ray binaries with steady nuclear burning on the WD surface, the dominating source of luminosity is the nuclear burning rather than the accretion luminosity. However, in WD systems exhibiting nova outbursts, which are relatively short-lived, the accretion luminosity should be the greatest source of luminosity during the quiescent stages, and is therefore still very important to consider.

The associated temperature

The continuum spectrum associated with the accretion luminosity can be characterized by a temperature T_{rad} , which is defined in such a way that the energy $h\nu$ of a “typical” photon emitted, is of the order of kT_{rad} . The blackbody temperature of the source if it radiated the accretion luminosity as a blackbody spectrum, is (see also §2.6.5)

$$T_{\text{b}} = \left(\frac{L_{\text{acc}}}{4\pi R_1^2 \sigma} \right)^{1/4} . \quad (2.27)$$

One can also define a temperature T_{th} that would be reached by the accreted material if *all* its gravitational potential energy was converted into thermal energy. The potential energy released for each proton-electron pair is

$$\frac{GM_1(m_{\text{p}} + m_{\text{e}})}{R_1} \approx \frac{GM_1 m_{\text{p}}}{R_1} , \quad (2.28)$$

where m_{p} is the mass of a proton and m_{e} the mass of an electron. Equating this to the thermal energy of both the protons and the electrons, $3kT$, where k is Boltzmann’s constant, yields

$$T_{\text{th}} = \frac{GM_1 m_{\text{p}}}{3kR_1} . \quad (2.29)$$

²The WD radius derived with Eq. (2.25) can be considered as a lower limit when considering hot WDs, as shown by Panei, Althaus and Benvenuto (2000).

For an optically thin accretion flow, $T_{\text{rad}} \sim T_{\text{th}}$. If the accretion flow is optically thick, one would have $T_{\text{rad}} \sim T_{\text{b}}$. In general, therefore, $T_{\text{b}} \lesssim T_{\text{rad}} \lesssim T_{\text{th}}$.

The Eddington limit

The radiation field produced by accretion of matter onto the surface of the primary will interact with the accreting gas through Thomson scattering. The Thomson scattering cross-section for protons is $\sim 3 \times 10^{-7}$ times smaller than that for electrons, so primarily the electrons will be scattered, but the electrostatic attraction between the electrons and protons will drag the protons along. This process imposes a significant limitation on the accretion process, resulting in the existence of a maximum accretion luminosity for a star with a given mass, which is known as the Eddington luminosity.

The Eddington limit represents the limit posed as a result of the inward gravitational force on each electron-proton pair equalling the outward radiation force on the pair. Assuming that the accreting stream is steady and spherically symmetric, and that the material consists mainly of ionized hydrogen, it can be shown that the corresponding luminosity is given by

$$L_{\text{E}} = \frac{4\pi c G M_1 m_{\text{p}}}{\sigma_{\text{T}}} \cong 1.25 \times 10^{38} \left(\frac{M_1}{M_{\odot}} \right) \text{ erg s}^{-1}, \quad (2.30)$$

where σ_{T} is the Thomson cross-section. For accretion exceeding L_{E} , the radiation pressure will inhibit the accretion process, “switching off” the source if all the source luminosity was derived from accretion. If, for example, part of the luminosity was derived from nuclear burning, super-Eddington accretion would cause the outer layers of accreted material to be blown off.

2.4 The accretion disc

The mass transfer process described in the previous section often causes the formation of a disc-like structure around the primary that is called an *accretion disc*. The discussion of accretion discs in this section is largely based on the approach of Frank et al. (2002, pp. 58-70, 80-98, 129-132), with occasional reference to other sources.

2.4.1 Accretion disc formation

The supersonic stream of material transferred through the L_1 point will have a substantial amount of specific angular momentum because of the orbital motion. At the onset of mass transfer, it will be deflected sideways due to the Coriolis effect, and swirl around the primary star (if the closest approach of the stream is larger than the radius of the star), following a ballistic trajectory that is determined by the “injection velocity” at L_1 and the gravitational field of the primary, and intersect

its own path, resulting in a rather messy impact and energy dissipation via shocks.

Up to this point, the accretion stream has not had a chance to rid itself of angular momentum, and through shocks and radiation it will tend to settle into an orbit around the primary in the binary plane with the lowest energy for its angular momentum, i.e. a circular Keplerian orbit. The radius of this particular orbit is called the *circularization radius*, which for Roche lobe overflow is given by the formula

$$\frac{R_{\text{circ}}}{a} = (1 + q) (0.500 - 0.227 \log q)^4 . \quad (2.31)$$

R_{circ} is smaller than the Roche lobe of the primary, and for a compact primary it is larger than the primary radius R_1 . The angular velocity of material in a Keplerian orbit is dependent on the radial distance from the centre of the primary according to

$$\Omega_K(R) = \left(\frac{GM_1}{R^3} \right)^{1/2} , \quad (2.32)$$

while the circular velocity is given by

$$v_K(R) = R \Omega_K(R) = \left(\frac{GM_1}{R} \right)^{1/2} . \quad (2.33)$$

The specific angular momentum is

$$R^2 \Omega_K(R) = (GM_1 R)^{1/2} . \quad (2.34)$$

Thus, material orbiting at a larger distance from the primary will have a smaller angular and circular velocity, but a higher specific angular momentum.

In the initial ring of material at R_{circ} , dissipative processes like collisions, shocks, and also viscous dissipation due to the velocity gradient between layers, will effectively cause the gravitational potential energy of the material to be converted to heat, some of which is radiated away. This loss of potential energy causes some of the gas to sink deeper into the gravitational potential well of the primary and therefore lose angular momentum. However, to conserve the total angular momentum, other material must now move out to larger orbits with a higher angular momentum, forming an extended structure called the accretion disc.

It is useful to utilize cylindrical polar coordinates when describing the disc, with R the radial distance from the centre of the primary, ϕ the azimuthal coordinate and z the height within the disc. In an accretion disc, material is spiralling inwards towards the primary, and angular momentum is transferred outwards. At the outer edge of the disc, tidal interactions with the secondary star are expected to feed the

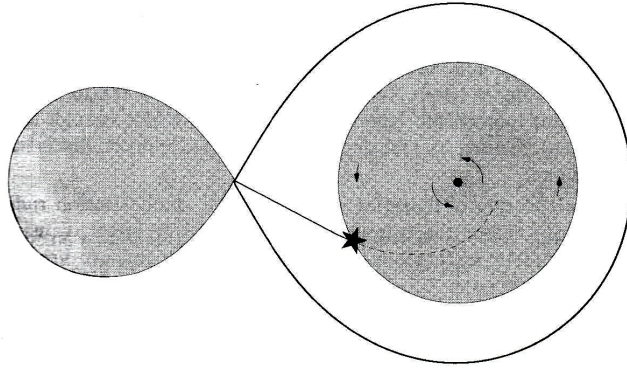


Fig. 2.2: Top view of an accreting binary with an accretion disc and a “bright spot” where the incident stream collides with the disc. (Adopted from Hellier, 2001, Fig. 2.8, p. 26.)

angular momentum back into the binary orbit. The outer disc radius R_{out} is therefore larger than R_{circ} , but can not be larger than the primary Roche lobe radius $R_{\text{L},1}$.

Once the accretion disc has formed, the stream of material from the L_1 point will impinge on the edge of the disc, as illustrated in Fig. 2.2. This turbulent encounter will cause the kinetic energy of the stream to be converted to heat and radiated away. This causes the formation of a so-called *bright spot*, which in some CVs emits $\sim 30\%$ of the total light of the system (e.g. Hellier, 2001, p. 27).

2.4.2 Disc viscosity

The differential rotation in the disc means that the adjacent annuli are sliding or shearing past each other, generating viscous stresses. The viscosity will try to force adjacent annuli into corotation by speeding up the outer annulus and slowing down the inner one. This enables transport of angular momentum radially, i.e. outwards through the disc. This is known as *shear viscosity*.

Molecular viscosity alone can not account for the effective angular momentum transfer taking place in accretion discs; in fact, it is several orders of magnitude too small. To account for this difficulty, it has been proposed that accretion discs are turbulent. The exchange of turbulent blobs with different amounts of angular momentum between neighbouring annuli will act as a viscosity mechanism by aiding the outward transport of angular momentum. The *origin* of the turbulence in the disc will be discussed in more detail in §2.7. Here only some key features will be highlighted.

The size of a turbulent cell in an accretion disc can not exceed the disc thickness H , and the turbulent turnover velocity is very unlikely to be supersonic, otherwise the turbulent motion would be broken down to thermal motion by shocks. Because the kinematic viscosity can be expressed as the product of the scale and speed of the motion causing the viscosity, the kinematic viscosity due to turbulence can be



Fig. 2.3: A schematic representation of a thin accretion disc. (Adopted from Biskamp, 2003, Fig. 11.3, p. 244.)

written as

$$\nu_{\text{kin}} = \alpha c_s H , \quad (2.35)$$

where c_s is the sound speed in the gas, which is $\sim 10 \text{ km s}^{-1}$. This is only a parametrization rather than a solution, as our uncertainty about the magnitude of turbulent viscosity is now just confined to the viscosity parameter α , which has an upper limit of 1 and is not to be confused with the mass loss fraction in §2.3.2. Eq. (2.35) represents the famous α -prescription of Shakura and Sunyaev (1973).

Standard theoretical models for accretion discs with Keplerian rotation make use of this so-called “alpha viscosity” parametrization, and usually assume what is known as a “thin disc”. These discs are commonly referred to as α -discs. In the thin disc approximation, the scaleheight H of the disc is much smaller than its radius R , with $H \propto R^{9/8}$ (e.g. Biskamp, 2003, p. 243). The azimuthal speed v_ϕ (equivalent to the Keplerian circular velocity) is highly supersonic (of the order of 1000 km s^{-1}), while the radial inward drift speed v_R is subsonic, typically $\sim 0.3 \text{ km s}^{-1}$. The total mass of the disc itself is negligible compared with that of the central compact object. The latter is generally a very realistic assumption due to the relatively low density of the disc (number density of the order $\sim 10^{15} \text{ cm}^{-3}$, e.g. Biskamp, 2003, p. 238).

The above-mentioned $H \propto R^{9/8}$ dependence means that the disc has a concave shape, flaring slightly at its edges (see Fig. 2.3). However, this is assuming that α is independent of R , which needs not be true. The disc will indeed still be concave if α increases with R more slowly than $\sim R^{5/4}$, but if α increases more rapidly with R , the disc may be convex in some regions.

2.4.3 Disc luminosity and spectrum

In a steady, thin, Keplerian disc, the viscous dissipation per unit surface area of the disc can be shown to be

$$D(R) = \frac{3GM_1\dot{m}_{\text{acc}}}{8\pi R^3} \left[1 - \left(\frac{R_1}{R} \right)^{1/2} \right] , \quad (2.36)$$

provided that the primary is a slow rotator (see §2.5) and thus does not disrupt the accretion flow at R . Using Eq. (2.36), it can be shown that the total luminosity of

the disc is

$$L_{\text{disc}} = \frac{GM_1 \dot{m}_{\text{acc}}}{2R_1} , \quad (2.37)$$

which is one half of the accretion luminosity, given in Eq. (2.24). The remaining half of the gravitational potential energy is converted to the orbital kinetic and turbulent energy of the disc material, and is only radiated close to the primary, as explained in §2.5.

Under the assumption that the accretion disc is optically thick in the z -direction, each disc element can be considered to radiate approximately as a blackbody with temperature $T_{\text{disc}}(R)$, i.e.

$$D(R) = \sigma T_{\text{disc}}^4(R) , \quad (2.38)$$

where σ is the Stefan-Boltzmann constant. From Eq. (2.36), this yields a radius-dependent disc temperature of

$$T_{\text{disc}}(R) = \left\{ \frac{3GM_1 \dot{m}_{\text{acc}}}{8\pi R^3 \sigma} \left[1 - \left(\frac{R_1}{R} \right)^{1/2} \right] \right\}^{1/4} . \quad (2.39)$$

For radii $R \gg R_1$ in the disc, the relation

$$T_{\text{disc}}(R) = T_* \left(\frac{R}{R_1} \right)^{-3/4} \quad (2.40)$$

holds, where

$$T_* = \left(\frac{3GM_1 \dot{m}_{\text{acc}}}{8\pi R_1^3 \sigma} \right)^{1/4} \quad (2.41)$$

is of the order of T_b defined in §2.3.3. Differentiation of Eq. (2.39) shows that the maximum value of $T_{\text{disc}}(R)$ is

$$T_{\text{disc}} \left(\frac{49R_1}{36} \right) = 0.488 T_* . \quad (2.42)$$

Neglecting the optically thin disc atmosphere for now, the spectrum emitted by each disc element can be approximated by the Planck function. The total flux at frequency ν observed from the disc by an observer at a distance D is then

$$F_\nu = \frac{4\pi h \nu^3 \cos i}{c^2 D^2} \int_{R_{\text{in}}}^{R_{\text{out}}} \frac{R dR}{e^{h\nu/kT_{\text{disc}}(R)} - 1} . \quad (2.43)$$

For non-magnetic WDs and neutron stars, R_{in} is equal to the stellar radius, while for magnetized objects, R_{in} is equal to the Alfvén radius (see §2.5).

The shape of the accretion disc spectrum for different $R_{\text{out}}/R_{\text{in}}$ ratios is shown in

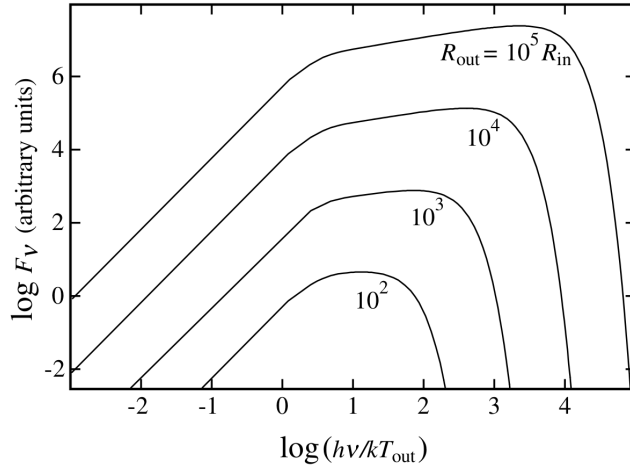


Fig. 2.4: Continuum spectra for steady, optically thick accretion discs with different $R_{\text{out}}/R_{\text{in}}$ ratios radiating locally as blackbodies. Note the $F_\nu \propto \nu^{1/3}$ dependence in the intermediate region. (Adopted from Frank et al., 2002, Fig. 5.2, p. 92.)

Fig. 2.4. For $\nu \ll kT(R_{\text{out}})/h$, the Planck function has the Rayleigh-Jeans form, i.e. $2kT\nu^2/c^2$, yielding $F_\nu \propto \nu^2$, while for $\nu \gg kT_*/h$, the Planck function has the Wien form, i.e. $2h\nu^3 c^{-2} e^{-h\nu/kT}$ and the spectrum has an exponential form. For frequencies in between, it can be shown that $F_\nu \propto \nu^{1/3}$.

The power law region of the spectrum in Fig. 2.4 with $F_\nu \propto \nu^{1/3}$ is considered to be characteristic of an accretion disc. However, this flattened portion of the spectrum is only substantial if the outer disc temperature is much lower than that of the inner disc, in other words if $R_{\text{out}} \gg R_{\text{in}}$. For WDs, $R_{\text{out}} \sim 10^2 R_{\text{in}}$, and the expected disc spectrum does not deviate greatly from an ordinary blackbody curve.

For an accreting WD in a binary system where the optical luminosity mostly originates from a standard accretion disc, the absolute visual magnitude³ is related to other system parameters by (Webbink et al., 1987)

$$M_V^{\text{disc}} \approx -9.48 - \frac{5}{3} \log \left(\frac{M_{\text{wd}}}{M_\odot} \frac{\dot{m}_{\text{acc}}}{M_\odot \text{ yr}^{-1}} \right) - \frac{5}{2} \log (2 \cos i) . \quad (2.44)$$

In addition to the characteristic continuum emission, spectral lines are also observed from accretion discs. Although absorption lines may be observed in some cases, accretion disc spectra are usually characterized by emission lines (e.g. Hellier, 2001, pp. 34-37). These emission lines originate in the hot chromosphere of the disc which forms due to dissipation from the surface. The heating of the disc by X-rays from the central object can cause the formation of a more extended disc atmosphere, called a *corona*. This means an extended optically thin region, with strong line emission.

³The absolute magnitude is the apparent magnitude that would be observed if the source was at a distance of 10 pc from the observer (e.g. Karttunen et al., 2007, p. 88).

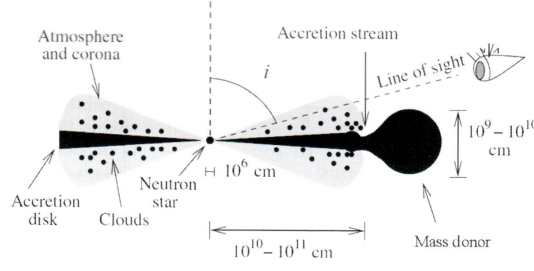


Fig. 2.5: A schematic representation of an accretion disc corona in an LMXB. It illustrates that, at high inclinations, the neutron star itself may be obscured, but the observer may still see X-rays scattered from the corona. (Adopted from Seward and Charles, 2010, Fig. 11.46, p. 208.)

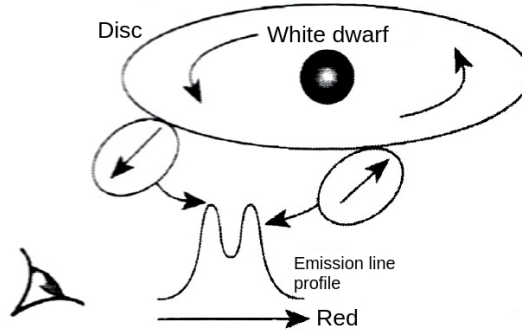


Fig. 2.6: The formation of a double-peaked emission line by an accretion disc viewed at high inclination. (Adopted from Seward and Charles, 2010, Fig. 10.3, p. 147.)

In Fig. 2.5, an accretion disc corona around a neutron star in an LMXB is illustrated. A similar effect can be expected in supersoft X-ray binaries, which have a very high X-ray luminosity. However, in supersoft X-ray binaries the corona temperature will be lower and the thickness smaller than in X-ray binaries, due to the lower effective temperature of the central X-ray source.

When an accretion disc is observed at a high inclination, i.e. when the plane of the disc is close to the line of sight, the emission lines have a characteristic double-peaked structure (Smak, 1969; Horne and Marsh, 1986). This is because the high-speed azimuthal rotation of the material in the disc causes a large distribution of Doppler shifts to be associated with the disc emission. Emission from disc regions receding from the observer is redshifted, with emission from approaching regions being blueshifted. This is illustrated in Fig. 2.6. These double-peaked emission lines are tell-tale signatures of the presence of an accretion disc and can be utilized by the observational astronomer as a diagnostic tool to unravel the mysteries of these sources.

2.4.4 Irradiated discs

In many accreting binaries, the accretion disc is irradiated significantly by the central primary star. In some cases, like in supersoft X-ray binaries, the disc luminosity due to irradiation can even exceed the intrinsic accretion luminosity of the disc. Rather than the characteristic $F_\nu \propto \nu^{1/3}$ shape of the spectrum of a standard accretion disc at intermediate frequencies, a relation of $F_\nu \propto \nu^{-1}$ is applicable to irradiated discs (see also Fukue and Matsumoto, 2001). It can be shown that the effective temperature T_{irr} of the disc caused by irradiation by an extended central source is given by

$$\left(\frac{T_{\text{irr}}}{T_{\text{eff}}}\right)^4 \approx \frac{2}{3\pi} \left(\frac{R_1}{R}\right)^3 (1 - \beta) \quad (2.45)$$

for $R \gg R_1$, where T_{eff} represents the temperature of the central source, and β the disc albedo, i.e. $(1 - \beta)$ is the fraction of incident radiation absorbed by the disc. For $R \rightarrow R_1$,

$$\left(\frac{T_{\text{irr}}}{T_{\text{eff}}}\right)^4 \rightarrow \frac{(1 - \beta)}{2}, \quad (2.46)$$

yielding a larger value than what would have been obtained with Eq. (2.45), but in both cases $T_{\text{irr}} < T_{\text{eff}}$ (see also Friedjung, 1985). The formulas above assume an optically thick disc. An optically thin disc with the same brightness would probably have a higher temperature. Thus in a thin disc, for $R \gg R_1$, the temperature $T_{\text{irr}} \propto R^{-3/4}$ has the same dependence on the radius as the effective temperature T_{disc} of a disc without irradiation. However, in an irradiated disc with a flaring thickness, $T_{\text{irr}} \propto R^{-3/7}$ for large R (Fukue, 2012 and references therein). Comparison with Eq. (2.40) shows that

$$\left(\frac{T_{\text{irr}}}{T_{\text{disc}}}\right)^4 = \frac{4}{9\pi} \frac{L_*}{L_{\text{acc}}} (1 - \beta), \quad (2.47)$$

for $R \gg R_1$, where $L_* = 4\pi R_1^2 \sigma T_{\text{eff}}^4$ represents the luminosity of the central extended source. Therefore, the temperature of the disc due to irradiation by an extended source with luminosity $L_* > 7.1 L_{\text{acc}} (1 - \beta)^{-1}$ will dominate the disc temperature due to viscous heating at all radii. Analytical studies by Fukue (2012) showed that in such cases, the emergent accretion disc spectra have two components: one multi-colour blackbody component representing the reprocessed radiation, with $T_{\text{irr}} < T_{\text{eff}}$, and one single blackbody with temperature equal to T_{eff} , representing scattering of the incident light. When irradiation heating dominates, the disc is often called a *passive disc*, while in an *active disc*, viscous heating dominates.

Pringle (1996) also showed that when an accretion disc is strongly irradiated by a central emission source, the disc can become unstable to warping. This occurs because of the torque that is exerted on the accretion disc by the absorbed radiation

that is re-emitted from the disc surface. A warped disc can be expected to “wobble”, or precess, and if an outflow/jet originates from the centre of the disc, the outflow will also precess (see Southwell, Livio and Pringle, 1997 and references therein).

According to Papaloizou and Pringle (1977), the disc in an accreting WD binary will be tidally truncated at $\sim 70\%$ of $R_{L,1}$, the radius of the Roche lobe of the WD. For warping to have an effect on an accretion disc, the warping should obviously occur within the extent of the accretion disc. Also, the warping instability is expected to occur only for disc radii satisfying the relation (Pringle, 1996; Livio and Pringle, 1996)

$$\frac{R_{\text{disc}}}{R_{\text{wd}}} \gtrsim \xi \eta^2 \left(\frac{L_{\text{nuc}}}{L_{\text{acc}}} \right)^{-2} \frac{R_{\text{wd}}}{2GM_{\text{wd}}/c^2}, \quad (2.48)$$

where L_{nuc} represents the bolometric luminosity, which is dominated by nuclear burning. The factor ξ is a constant. The parameter η represents the ratio of the (R, z) and (R, ϕ) viscosities in the disc. Adopting the analytic estimate of $\xi \sim 32\pi^2$, and assuming $\eta \sim 1$, implies the following upper limit in the accretion rate for warping to occur:

$$\dot{m}_{\text{acc}} \lesssim \frac{1}{4\pi c G^{1/2}} \left(\frac{R_{\text{disc}}}{R_{\text{wd}}} \right)^{1/2} L_{\text{nuc}} \left(\frac{R_{\text{wd}}}{M_{\text{wd}}} \right)^{1/2}. \quad (2.49)$$

Accretion rates higher than this will stabilize the disc against warping. Considering typical values of the parameters in supersoft X-ray binaries, Eq. (2.49) can be rewritten as

$$\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}} \lesssim 0.81 \left(\frac{R_{\text{disc}}}{100 R_{\text{wd}}} \right)^{1/2} \left(\frac{L_{\text{nuc}}}{10^{38} \text{ erg s}^{-1}} \right) \left(\frac{R_{\text{wd}}}{4.9 \times 10^8 \text{ cm}} \right)^{1/2} \left(\frac{M_{\text{wd}}}{\text{M}_{\odot}} \right)^{-1/2}. \quad (2.50)$$

For mass accretion rates satisfying this limit, the accretion disc will probably be unstable to the warping instability.

2.4.5 Properties of accretion discs in supersoft X-ray binaries

In SSSs, the nuclear burning luminosity is much larger than the accretion luminosity. Therefore, the emission from the accretion disc consists primarily of reprocessed radiation from the central X-ray source. Popham and Di Stefano (1996) showed that the observed optical and UV flux from several SSSs can be explained very well by a combination of the following: the WD spectrum, the accretion disc with the reprocessing of radiation included, as well as disc flaring, and also some emission of reprocessed radiation from the heated side of the donor star.

According to Schandl, Meyer-Hofmeister and Meyer (1997) and Meyer-Hofmeister,

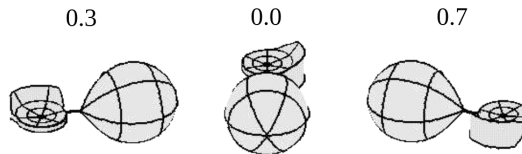


Fig. 2.7: The accretion disc rim in an SSS at orbital phases of 0.3, 0.0 and 0.7. (Adopted from Meyer-Hofmeister et al., 1997, Fig. 4.)

Schandl and Meyer (1997), the high accretion rate in SSSs may produce an optically thick rim with a height up to $z/R = 0.4$ that extends halfway around the accretion disc (Fig. 2.7). This is because the mass transfer stream incident on the disc at the bright spot causes a “bulge” or “spray” on the disc rim that consists of many individual gas blobs embedded in the surrounding corona. The blobs are bouncing up and down at the rim, and up to a certain height this “curtain” of blobs is optically thick. This extended “screen” then presents a larger cross-section to the primary star, and hence the reprocessed radiation from the disc rim makes it a more prominent contributor to the optical brightness relative to the other contributors. Meyer-Hofmeister et al. (1997) also suggested that the erratic long-term variability in some SSSs may be related to the changing extent and height of this “screen”.

Suleimanov, Meyer and Meyer-Hofmeister (1999) evaluated the reprocessing efficiency of soft X-rays by the disc by using irradiated stellar atmosphere modelling. It first appeared that the reprocessing of X-ray radiation is not effective enough to explain the high optical fluxes. However, Suleimanov, Meyer and Meyer-Hofmeister (2003) showed that adequate reprocessing of X-rays to optical light can be sustained during multiple scatterings between the gas blobs in the extended accretion disc rim.

The term “superdiscs” was also suggested to describe the accretion discs in SSSs (Fukue and Matsumoto, 2001), and refer to supercritical discs accreting at rates similar to or higher than the Eddington rate. Kitabatake, Fukue and Matsumoto (2002) proposed a scenario of highly super-Eddington accretion rates ($\sim 10^{-5} M_{\odot} \text{ yr}^{-1}$) at the edge of the superdisc, with most of the material expelled by a so-called superwind, yielding an accretion rate onto the WD of approximately 1 or 2 orders of magnitude lower than the mass transfer rate.

2.5 Accretion onto the white dwarf

After releasing one half of the available accretion energy while passing through the accretion disc, the transferred material still has to release the other half close to the WD before undergoing accretion. Depending on the strength of the WD magnetic field, the latter process can take place by one of two processes. If the WD magnetic field is weak, the disc extends down to the WD surface, and accretion takes place in

what is known as a boundary layer. On the other hand, if the WD has a substantial magnetic field, it can disrupt the inner parts of the disc and redirect the accretion flow onto the polecaps. In this section, the most relevant properties of these two mechanisms are summarized (see e.g. Frank et al., 2002, pp. 152-191 for a detailed discussion).

2.5.1 Boundary layer accretion

The WD itself will have a certain rotation period P_* . To prevent break-up, the angular velocity of the WD rotation should be smaller than the Keplerian velocity of the material at its equator, i.e.

$$\Omega_* < \Omega_K(R_1) . \quad (2.51)$$

In a Keplerian disc, $\partial\Omega/\partial R < 0$, therefore the angular velocity increases inwards, but for accretion to occur it needs to decrease to Ω_* at R_1 . The radius at which the decrease commences is $R = R_1 + b$, where we have $\partial\Omega/\partial R = 0$. The *boundary layer* refers to the region where $R_1 \lesssim R \lesssim R_1 + b$. It can be shown that the boundary layer is very thin for thin accretion discs, with

$$b \sim \frac{H^2}{R_1} \ll H \ll R_1 , \quad (2.52)$$

where H now refers the disc scaleheight measured inside the boundary layer. However, if it happens that the size of b is comparable to R_1 , the thin disc approximation will no longer be valid inside $R = R_1 + b$, and instead of a boundary layer, some sort of a thick disc (torus) structure will be formed.

A schematic diagram of a thin disc boundary layer is shown in Fig. 2.8. The energy emitted in the boundary layer will emerge through a region with a radial extent of $\sim H$ on the two faces of the accretion disc. If the boundary layer is optically thick, it will radiate roughly as a blackbody with a characteristic temperature

$$T_{\text{BL}} \sim T_* \left(\frac{T_{\text{S}}}{T_*} \right)^{1/8} , \quad (2.53)$$

where T_* is given by Eq. (2.41), and

$$T_{\text{S}} = \frac{3GM_1\mu m_{\text{H}}}{8kR_1} \quad (2.54)$$

is the shock temperature one would have obtained for radial accretion onto the WD. Here μ represents the mean molecular weight, and m_{H} the mass of a hydrogen atom. Comparison of the above with Eq. (2.42) shows that T_{BL} is higher than the temperature of the accretion disc, with the tail of the blackbody extending into the

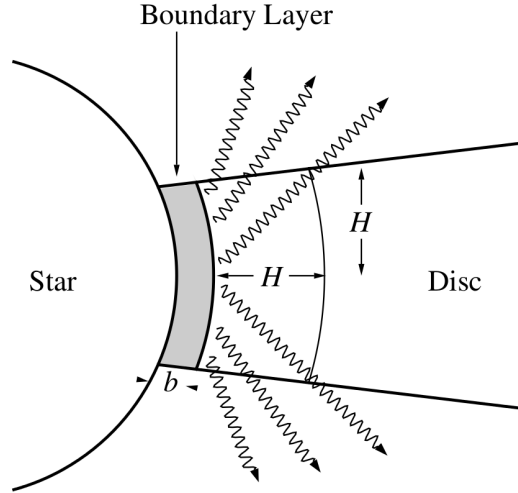


Fig. 2.8: A schematic diagram of an optically thick boundary layer in a plane with constant ϕ (not to scale). (Adopted from Frank et al., 2002, Fig. 6.2, p. 156.)

soft X-ray region. Hard X-ray emission from accreting white dwarfs with boundary layer accretion is thought to originate in an optically thin boundary layer, with shocks or turbulent viscous heating producing the hard X-ray temperatures.

2.5.2 Accretion onto a magnetized white dwarf

White dwarfs often possess magnetic fields of the order $B_1 \lesssim 10^8$ G. For a magnetic field with a dipole nature, the field strength varies with radial distance as $B \sim B_1(R_1^3/R^3)$. Let us first consider the simple case of quasi-spherical accretion flow. Closer than a certain distance from the WD, the magnetic pressure will dominate the ram and gas pressure of the accreting material. This distance is known as the *Alfvén radius*, and for the spherical flow it is given by

$$r_M = \left(\frac{B_1^4 R_1^{12}}{8GM_1 \dot{m}_{\text{acc}}^2} \right)^{1/7}. \quad (2.55)$$

Returning to disc accretion, the cylindrical Alfvén radius R_M (also known as the magnetospheric radius) must be found where the magnetic torque on the disc due to B_1 equals the viscous torque. It is found that R_M is of the same order as r_M : typically $R_M \sim 0.5r_M$ (although R_M may be up to 4 times larger than this). With this estimate, we can write

$$R_M \sim 0.5 \left(\frac{B_1^4 R_1^{12}}{8GM_1 \dot{m}_{\text{acc}}^2} \right)^{1/7}, \quad (2.56)$$

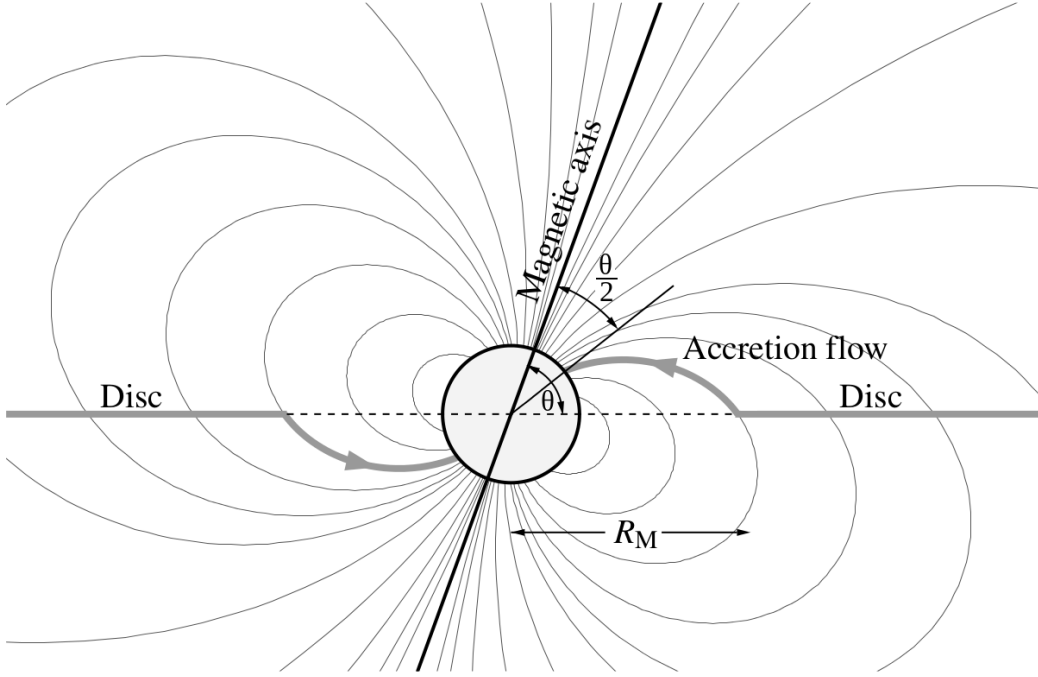


Fig. 2.9: A schematic diagram of accretion onto a magnetized WD, with material at a distance R_M from the WD being channelled towards the polecaps by the magnetic field. θ is the angle between the polar axis and the disc. (Adopted from Frank et al., 2002, Fig. 6.4, p. 161.)

or

$$R_M \sim 2.3 \times 10^9 \left(\frac{B_1}{10^7 \text{ G}} \right)^{4/7} \left(\frac{R_1}{4.9 \times 10^8 \text{ cm}} \right)^{12/7} \times \left(\frac{M_1}{M_\odot} \right)^{-1/7} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} M_\odot \text{ yr}^{-1}} \right)^{-2/7} \text{ cm} . \quad (2.57)$$

At the radius R_M , the accretion disc structure is disrupted by the magnetic field, and for $R < R_M$ the material will flow along the magnetic field lines, and be channelled onto the polecaps of the WD, as shown in Fig. 2.9.

A CV containing an accretion disc that is disrupted at $R_M > R_{\text{wd}}$ is called an *intermediate polar*. Intermediate polars exhibit coherent WD rotation periods that are shorter than the orbital period. In some CVs with very strong magnetic fields and short orbital periods, $r_M > a$, no disc is formed and the accreting gas couples directly with the magnetic field. These are known as AM Her systems, or *polars*, and the rotation period of the WD is tidally locked to the orbital period, or very nearly so.

Let us consider a supersoft X-ray binary containing a WD with $M_1 = 1 M_\odot$ and $R_1 = 4.9 \times 10^8 \text{ cm}$. In these sources, the accretion rate is typically of the order of

$10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}$, and a magnetic field strength of 10^7 G yields $R_M \sim 2.3 \times 10^9 \text{ cm} \sim 4.7 R_1$. This means that the type of scenario as in Fig. 2.9 will arise, with the disc structure being disrupted at a significant distance from the WD. On the other hand, for $B_1 < 6.6 \times 10^5 \text{ G}$, boundary layer accretion will occur, because $R_M < R_1$.

From the geometry in Fig. 2.9, one can see that accretion onto a polecap is limited to a region with half-angle $\theta/2$. It can easily be shown that the area of one accreting polecap as a fraction of the total stellar surface is

$$f \sim \frac{R_1 \sin^2 \theta}{4R_M} . \quad (2.58)$$

If accretion takes place on both polecaps, the fraction of the stellar surface that is actually accreting can be approximated by

$$f \sim \frac{R_1}{2R_M} \quad (2.59)$$

if $\theta \sim 90^\circ$. Neglecting the effect of plasma instabilities yields values between 10^{-1} and 10^{-4} for f .

As the accreting gas is channelled onto a polecap, a structure called an *accretion column* is formed, and it is in this column that the remaining half of the accretion luminosity is released. The material channelled into the column is highly supersonic, virtually in free-fall, and before accretion onto the WD surface occurs, a dramatic deceleration needs to take place. This leads one to expect the formation of a strong shock in the accretion column. Because the electron mass is so small, essentially all the kinetic energy in the flow is carried by the ions.

For all reasonable situations involving a single accretion column, it can be shown that the ions are stopped above the WD photosphere. In this strong shock region, the ions are decelerated to subsonic velocities, with most of the ordered kinetic energy of the supersonic incident flow converted to heat by dissipative processes. For an adiabatic shock, the post-shock velocity is smaller than the initial velocity by a factor of 4. The shock temperature is given by Eq. (2.54), which can be rewritten as

$$T_S = 7.6 \times 10^8 \left(\frac{M_1}{\text{M}_{\odot}} \right) \left(\frac{R_1}{4.9 \times 10^8 \text{ cm}} \right)^{-1} \text{ K} \quad (2.60)$$

with $\mu = 0.62$ for a fully ionized gas mixture with mass fractions of $X \sim 0.70$ for hydrogen, $Y \sim 0.28$ for helium and $Z \sim 0.02$ for metals. This shock-heated gas must now be cooled, and the high temperature suggests bremsstrahlung to be the dominant cooling mechanism, with the associated emission of hard X-rays. The geometry of the accretion column will determine the absorption effect of the hard

X-rays that will occur, and therefore the direction in which hard X-rays will be able to escape. If the accretion column is “tall and thin”, the cooling region below the shock front will be optically thick in the radial direction and optically thin in the “horizontal” direction. If the column is “short and fat”, the optical thickness in the radial direction may be smaller than that in the “horizontal” direction.

It must, however, be kept in mind that a substantial fraction of the radiation from the region below the shock will be radiated or transported to the photosphere of the WD at the base of the column. Here the hard X-rays will be absorbed and re-emitted, with the re-emitted spectrum most probably resembling a blackbody with temperature

$$T_b = \left(\frac{L_{\text{acc}}}{4\pi R_1^2 f \sigma} \right)^{1/4} \quad (2.61)$$

$$\sim 10^5 \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}} \right)^{1/4} \left(\frac{M_1}{\text{M}_{\odot}} \right)^{1/4} \left(\frac{R_1}{4.9 \times 10^8 \text{ cm}} \right)^{-3/4} \left(\frac{f}{10^{-2}} \right)^{1/4} \text{ K} , \quad (2.62)$$

showing that the blackbody radiation will consist mainly of soft X-rays. This brings us to another potential cooling mechanism: the electrons can interact with the soft photons through inverse Compton scattering, during which the photons are upscattered to higher energies, while the electron gas is cooled. Cyclotron emission will also be a cooling factor in the presence of a magnetic field, but will in general be more of an important signature process than a major cooling mechanism. It will only be for very low $\dot{m}_{\text{acc}}$ or very large B_1 that cyclotron emission will make out a significant fraction of the total accretion luminosity.

Theoretical considerations of this accretion column model leads to the expectation that the largest fraction of the radiation should be in the form of hard X-rays. However, observations of such systems reveal the contrary, with the spectrum dominated by soft X-ray emission and with the hard X-ray luminosity often only 10% of the system luminosity. In order to account for this discrepancy, it has been suggested that the accretion column model should be slightly modified to describe polecap accretion not in terms of a single large accretion column, but in terms of small accreting blobs, each creating their own small accretion column (Frank, King and Lasota, 1988). In this revised model, the original flow is broken up into well-separated blobs above the WD surface. Diffuse blobs will be shocked above the WD surface, while dense blobs can penetrate into the photosphere before being shocked, enhancing the soft X-ray production (see Fig. 2.10).

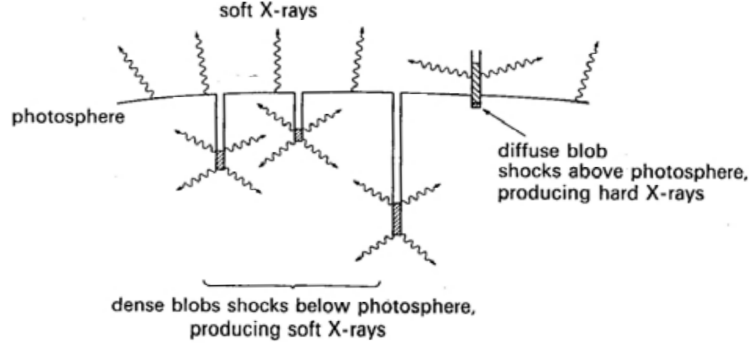


Fig. 2.10: Accretion of multiple blobs onto the polecap of a magnetic WD. (Adopted from Frank et al., 2002, Fig. 6.15, p. 190.)

2.6 The white dwarf primary

2.6.1 White dwarf rotation

As mentioned in §2.5.1, the WD will be rotating around its own axis with angular velocity Ω_* , and an associated period $P_* = 2\pi/\Omega_*$. Generally, the axis of rotation is perpendicular to the disc, and the rotation in the same azimuthal direction as that of the disc. For the WD to remain a coherent whole, Ω_* can not be larger than the Keplerian angular velocity $\Omega_K(R_1)$ at the surface of the WD at its equator, the so-called *break-up* angular velocity. This places a lower limit on the spin period, which is ~ 6 s for a solar-mass WD.

In the case of a magnetized WD, its magnetic field pattern (the *magnetosphere*) will be rotating together with the WD. The so-called *fastness parameter*, given by

$$\omega_* = \frac{\Omega_*}{\Omega_K(R_M)} \quad (2.63)$$

can be used to compare the magnitude of Ω_* to the Keplerian angular velocity at the Alfvén radius (e.g. Frank et al., 2002, p. 160). If $\omega_* < 1$, the primary is a “slow rotator”, while $\omega_* \sim 1$ for a “fast rotator”. A related concept is that of the *corotation radius*, which is the radial distance from the primary where the local Keplerian angular velocity equals Ω_* . The corotation radius is therefore given by (e.g. Frank et al., 2002, p. 163)

$$R_\Omega = \left(\frac{GM_1 P_*^2}{4\pi^2} \right)^{1/3} = 1.5 \times 10^8 P_*^{2/3} \left(\frac{M_1}{M_\odot} \right)^{1/3} \quad (2.64)$$

The condition $\omega_* < 1$ for a slow rotator can therefore also be expressed as $R_\Omega > R_M$. This is especially important when considering magnetospheric outflows, as will be discussed in §2.8.2.

If the spin period is to be observed, there needs to be some sort of asymmetry in the emission from the WD surface. For a magnetized WD, this is caused by a misalignment of the rotation and magnetic axes of the accreting WD, resulting in pulsed emission where the accretion heated pole (see Fig. 2.9) rotates through our line of sight to the system. The X-ray (and sometimes optical) emission from these accretion driven systems reveals the WD rotation period, which is typically of the order of tens to thousands of seconds.

Examples of WD spin periods are the rapid rotators like WZ Sge (28.96 s), AE Aqr (33.08 s), and intermediate polars like DQ Her (142 s), FO Aqr (1254 s) and TV Col (1910 s) (e.g. Norton, Wynn and Somerscales, 2004; Mukai, 2014). WD spin periods have also been observed in systems with surface hydrogen burning: the supersoft source XMMU J004252.5+411540 in M31 has exhibited a stable ~ 217.7 s supersoft X-ray modulation over a period of 5 years that is ascribed to the spin period of a magnetic WD (Trudolyubov and Priedhorsky, 2008). Similarly, the nova M31N 2007-12b exhibited a 1110 s stable X-ray periodicity during its supersoft phase, which is interpreted as the spin period of the WD in an intermediate polar system (Pietsch et al., 2011). Osborne et al. (2001) detected a ~ 865 s supersoft oscillation from the transient source XMMU J004319.4+411759, which is also interpreted as the spin period of an accreting WD.

During the supersoft phase of its 2006 eruption, the recurrent nova RS Oph was reported to exhibit a multiperiodic X-ray modulation of 35 s that became increasingly coherent (see Osborne et al., 2011 and references therein). These authors initially ascribed the modulation to non-radial pulsations driven by nuclear burning instabilities. However, it is argued by Orio and collaborators (2013, personal communication) that the 35 s period, now stabilized, is actually the WD spin period, and that the initial variability in the period was due to it being observed from a precessing jet originating at the polar caps of the WD. Similar periodicities were observed during the supersoft phase of the 2009 eruption of KT Eri (35 s) (Beardmore et al., 2010), the supersoft phase of the 2013 eruption of V339 Del (54 s) (Beardmore et al., 2013), and a ~ 163 s period in 2014 during the supersoft phase of the outburst of V745 Sco (Beardmore et al., 2014). More information can also be found in Ness et al. (2015), who also reported on a variable ~ 33 s period in nova LMC 2009a. A coherent 56.825 s spin period was discovered in V842 Cen (Woudt, Warner, Osborne and Page, 2009), but the latter periodicity was detected in photometric data several decades after the nova outburst.

Just like the position of a spectral line from a star in a binary system will oscillate about a central position due to the Doppler effect, the frequency of the observed spin period will also undergo cyclic modulations as the WD moves through its binary

orbit. If the orbital ephemeris is known, this contribution can be removed from the signal; or, conversely, the orbital period can be determined from such a pulse-timing analysis. From the Doppler formula in Eq. (2.2), it can be shown that the maximum shift in the observed spin period is given in terms of the radial velocity semi-amplitude of the WD orbital motion by

$$\frac{\Delta P_*}{P_*} = \left(\frac{K_1}{c} \right) \quad (2.65)$$

$$\sim 0.0003 \left(\frac{K_1}{100 \text{ km s}^{-1}} \right) , \quad (2.66)$$

assuming a circular binary orbit.

In many binaries, particularly with a neutron star primary, gradual changes in the primary spin period are observed which are not related to the binary motion described above. A steady decrease in P_* (spin-up) is presumably caused by the torques exerted by the accretion disc on the primary, and it is thought that the occasional increases in P_* (spin-down) may be caused by accretion torque modulations or possible changes in the internal neutron star structure. Similar changes in P_* have been observed in WDs, but these changes occur much slower because of the much higher moment of inertia of a WD, compared to accreting neutron stars. For a slow rotator, one can write the following equation for the rate of change of the spin angular velocity due to the accretion of angular momentum from an accretion disc (e.g. Frank et al., 2002, pp. 162-163):

$$\dot{\Omega}_* = \dot{m}_{\text{acc}} (GM_1 R_M)^{1/2} I^{-1} , \quad (2.67)$$

where I is the moment of inertia of the primary. Utilizing Eq. (2.56) and $I \sim M_1 R_1^2$, this can be rewritten instructively as the rate of change of the spin period,

$$\begin{aligned} \dot{P}_* = & 1.2 \times 10^{-9} \left(\frac{P_*}{1000 \text{ s}} \right)^2 \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_\odot \text{ yr}^{-1}} \right)^{6/7} \left(\frac{M_1}{\text{M}_\odot} \right)^{-4/7} \times \\ & \left(\frac{R_1}{4.9 \times 10^8 \text{ cm}} \right)^{-8/7} \left(\frac{B_1}{10^7 \text{ G}} \right)^{2/7} \text{ s s}^{-1} . \end{aligned} \quad (2.68)$$

Let us again consider a solar-mass WD in a supersoft X-ray binary accreting from a disc at $\sim 10^{-7} \text{ M}_\odot \text{ yr}^{-1}$. If the spin period of this WD is currently 1000 s, and it has a strong magnetic field of 10^7 G , the equation above shows that $\dot{P}_* \sim 1.2 \times 10^{-9} \text{ s s}^{-1}$, implying that it will take approximately 26 years for the WD to be spun up with 1 s to $P_* = 999 \text{ s}$. It is clear that a weaker magnetic field will mean a lower spin-up rate, as will a lower accretion rate. Also, the shorter the spin period becomes, the lower the spin-up rate will be.

In principle, any accreting primary can be spun up until it becomes a fast rotator, i.e. $\omega_* \sim 1$. In such systems, effects like interaction between the magnetic field and the disc outside the Alfvén radius, as well as the loss of angular momentum when material is ejected along the magnetic field lines will also play a role. In fact, the spin period should reach an equilibrium value when $\omega_* \sim 1$ and $R_\Omega \sim R_M$, which becomes

$$P_{\text{eq}} = 61 \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_\odot \text{ yr}^{-1}} \right)^{-3/7} \left(\frac{M_1}{\text{M}_\odot} \right)^{-5/7} \left(\frac{R_1}{4.9 \times 10^8 \text{ cm}} \right)^{18/7} \left(\frac{B_1}{10^7 \text{ G}} \right)^{6/7} \text{ s} \quad (2.69)$$

from Eq. (2.56) and Eq. (2.64).

2.6.2 Non-radial white dwarf pulsations

Apart from their rotation, which may introduce periodicity into the lightcurves of SSSs, WDs are also known to pulsate. During the course of their lifetime, WDs pass through several instability strips where they become non-radial pulsating stars (see the review of Córscico, 2009). These are pulsations in the form of radius changes of the star, where “non-radial” refers to the type where some parts of the stellar surface move inwards while other parts move outwards, i.e. where spherical symmetry is not preserved. Non-radial pulsations are divided into spheroidal and toroidal modes, with the spheroidal modes of special interest for WDs. The latter modes can be further subdivided according to the nature of the main restoring force during the pulsations: gravity and buoyancy for the g - and f -modes, and pressure gradients for the p -modes.

Non-radial stellar pulsations are very complex phenomena, especially in the case of rotating stars, and a detailed discussion can be found in Maeder (2009, pp. 401-448). In general terms, a linear non-radial pulsation mode in a spherically symmetric star can be described as a three-dimensional standing wave in spherical coordinates (e.g. Córscico, 2009):

$$\psi'_{klm}(r, \theta, \phi, t) = \psi'_{klm}(r) Y_l^m(\theta, \phi) e^{i\sigma_{klm}t}, \quad (2.70)$$

where the prime indicates a small Eulerian perturbation of a quantity ψ , which can for example be the pressure or gravitational potential. The factor $\psi'_{klm}(r)$ represents the radial part of the eigenfunctions. $Y_l^m(\theta, \phi)$ are the spherical harmonics, and the factor $e^{i\sigma_{klm}t}$ represents the time variability as determined by the eigenfrequencies σ_{klm} . In the adiabatic approximation, the eigenvalues and eigenfunctions are obtained by solving a 4th order eigenvalue problem.

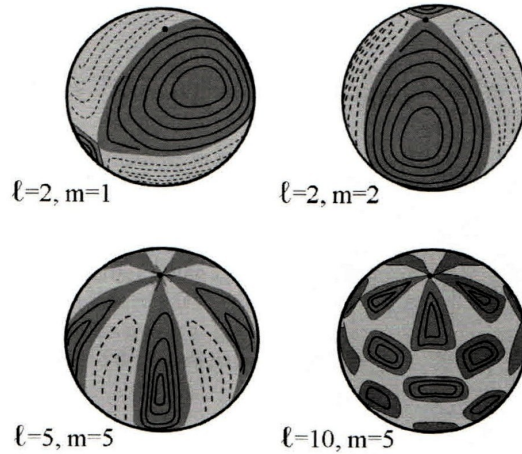


Fig. 2.11: Examples of some acoustic modes for a solar-type star, with different spherical harmonic degrees l and azimuthal numbers m . The lighter areas are moving towards the centre of the star, and the darker areas away from the centre. (Adopted from Maeder, 2009, Fig. 16.4, p. 418.)

The spherical harmonic degree $l = 0, 1, 2, \dots$ represents the number of nodal lines on the stellar surface, and the azimuthal number $m = 0, \pm 1, \dots, \pm l$ gives the number of nodal lines crossing the equator. When $l = 0$, the pulsations are radial. Examples of different combinations of spherical modes are given in Fig. 2.11. The radial order $k = 0, 1, 2, \dots$ indicates the number of nodes in the radial component of the eigenfunction, i.e. the number of concentric spherical surfaces with zero radial displacement (excluding the centre and the surface).

Many WDs and pre-WD stars exhibit pulsations with periods in the range of about a minute to more than an hour (e.g. Córscico, 2009). These oscillations are believed to be non-radial g -mode pulsations. Only modes with $l = 1$ and $l = 2$ have been identified at a high confidence level in WDs. Modes with higher l are thought to exist, but will be very difficult to detect, as the rapid modulations are expected to be washed out when observing the integrated light from the WD surface (Dziembowski, 1977). Theoretical stability analyses have also found unstable radial modes and p -modes with periods in the 0.1-10 s range (Saio and Wheeler, 1983; Kawaler, 1993), but these have not yet been observed in any WD.

Non-radial g -mode pulsations were first observed in the optical and UV emission from single WDs, with temperatures ranging from ~ 10000 to ~ 180000 K (Córscico, 2009 and references therein). Three distinct families of distinct pulsating white dwarfs are usually discriminated (see also the review of Fontaine and Brassard, 2008): the ZZ Ceti stars, with hydrogen atmospheres and $T \sim 12000$ K, the V777 Her stars with He atmospheres and $T \sim 25000$ K, and the hotter GW Vir stars (pre-WDs during the post planetary nebula phase) with He-C-O atmospheres and $T \sim 120000$ K.

These pulsations are often multi-periodic, and can sometimes be observed at about the same period(s) on time-scales of several years. The actual mechanism causing the pulsations is traditionally considered to be the κ - γ mechanism (e.g. Winget et al., 1982 and references therein). This mechanism involves the heating of a compressed gas consisting of partially ionized hydrogen and helium, due to the simultaneously increased metal opacities. Brickhill (1991) suggested convective driving to be another possible mechanism.

However, non-radial pulsations have also been observed in dwarf novae (Warner and Woudt, 2004, 2005), and more recently also in the X-ray lightcurves of hot accreting WDs exhibiting hydrogen shell burning, i.e. novae and SSSs, for example: Drake et al. (2003) found a strong periodic signal of ~ 41.7 min in the *Chandra* lightcurve of the nova V1494 Aql, together with a suite of periods from 8.8 to 57.7 min. An X-ray pulsation of ~ 22.1 min was discovered in the nova V4743 Sgr, with two weaker harmonic overtone periods at 11.1 and 7.5 min (Ness et al., 2003, see also Leibowitz et al., 2006). Another 38.4 min X-ray pulsation found in CAL 83 may be of a similar origin (Schmidtke and Cowley, 2006). In the presence of hydrogen burning, another possible mechanism for driving non-radial pulsations should be considered, namely the ϵ -mechanism, caused by the strong sensitivity of nuclear burning to temperature (Córscico and Althaus, 2014 and references therein).

2.6.3 Hydrogen burning on the surface of a white dwarf

Nuclear fusion can be ignited in an envelope of accreted hydrogen-rich material on the surface of the WD if a critical envelope mass $m_{\text{env}}^{\text{crit}}$ is reached, in which the temperature and pressure conditions required for the nuclear burning of hydrogen to helium by the CNO cycle can be sustained, i.e. $T \sim 10^8$ K and $P \gtrsim 10^{18}$ - 10^{20} dyn cm $^{-2}$ (10^{17} - 10^{19} N m $^{-2}$) (Fujimoto, 1982). The mass $\Delta m_{\text{env}}^{\text{crit}}$ is given by

$$\log \left(\frac{m_{\text{env}}^{\text{crit}}}{M_{\odot}} \right) \approx -2.862 + 1.542 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right)^{-1.436} \ln \left(1.429 - \frac{M_{\text{wd}}}{M_{\odot}} \right) - 0.197 \left[\log \left(\frac{\dot{m}_{\text{acc}}}{M_{\odot} \text{ yr}^{-1}} \right) + 10 \right]^{1.484} \quad (2.71)$$

(see also Prialnik and Kovetz, 1995; Townsley and Bildsten, 2004). The higher the WD mass and/or the accretion rate, the lower the required value of $m_{\text{env}}^{\text{crit}}$. For a $1 M_{\odot}$ WD accreting at $3 \times 10^{-7} M_{\odot} \text{ yr}^{-1}$ (approximately midway between the limits $\dot{m}_{\text{std}}$ and $\dot{m}_{\text{cr}}$ discussed below, for a hydrogen mass fraction of $X = 0.7$), we obtain $m_{\text{env}}^{\text{crit}} \sim 4 \times 10^{-6} M_{\odot}$.

As explained by Nomoto (1982) and refined by e.g. Hachisu and Kato (2001), the response of a WD to accretion can be divided into several different regimes. For

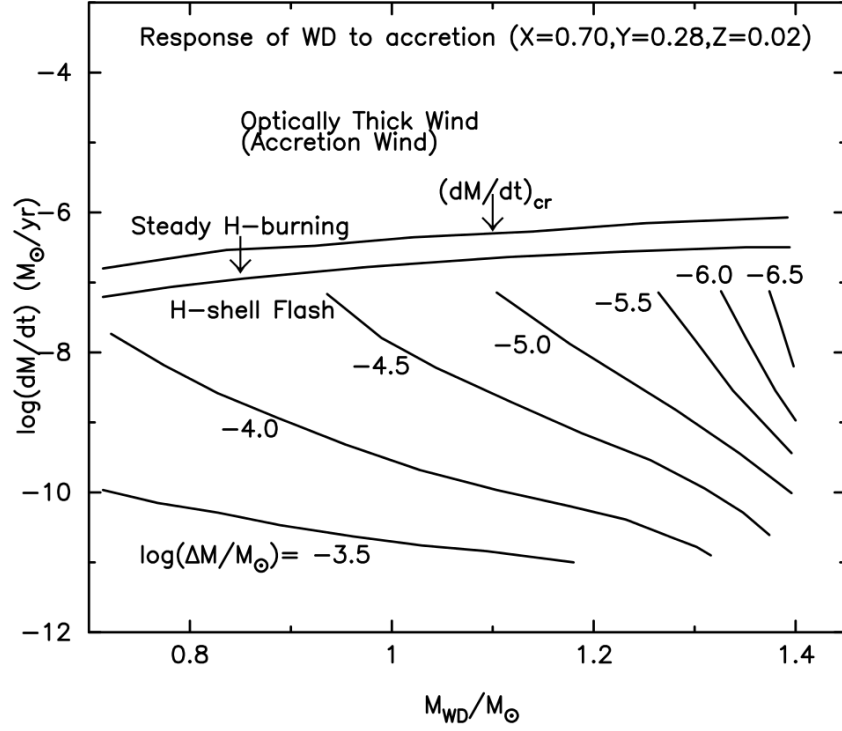


Fig. 2.12: The response of a WD to different accretion rates. The accretion rate dM/dt on the vertical axis is referred to in the text as $\dot{m}_{\text{acc}}$, while $(dM/dt)_{\text{cr}}$ represents $\dot{m}_{\text{cr}}$. ΔM refers to the critical envelope mass $m_{\text{env}}^{\text{crit}}$. (Adopted from Hachisu and Kato, 2001, Fig. 2, as adapted from Nomoto, 1982, Fig. 9.)

a WD with a given mass and accreting material with a certain composition, there exists a certain range of accretion rates for which steady hydrogen burning on the WD surface is possible (see Fig. 2.12). The composition of matter can be expressed in terms of the mass fractions X of H, Y of He, and Z of heavier elements (“metals”). For material with $Z = 0.02$ (i.e. solar metallicity), the lower limit of the steady burning regime is given by (Hachisu and Kato, 2001)

$$\dot{m}_{\text{std}} = 2.6 \times 10^{-7} \left(\frac{1.7 - X}{X} \right) \left(\frac{M_{\text{wd}}}{M_{\odot}} - 0.40 \right) M_{\odot} \text{ yr}^{-1}, \quad (2.72)$$

and the upper limit by

$$\dot{m}_{\text{cr}} = 5.3 \times 10^{-7} \left(\frac{1.7 - X}{X} \right) \left(\frac{M_{\text{wd}}}{M_{\odot}} - 0.40 \right) M_{\odot} \text{ yr}^{-1}. \quad (2.73)$$

For a WD in the steady burning regime, the luminosity L_{H} derived from the nuclear burning of hydrogen can be related to the accretion rate by (Iben, 1982)

$$\dot{m}_{\text{acc}} \approx \left(\frac{2.6 \times 10^{-8}}{X} \right) \left(\frac{L_{\text{H}}}{10^{37} \text{ erg s}^{-1}} \right) M_{\odot} \text{ yr}^{-1}. \quad (2.74)$$

Although hydrogen burning will be the dominant nuclear burning process, He burn-

ing will also take place, both contributing to the total nuclear burning luminosity L_{nuc} . Adopting $X = 0.64$ and $Y = 0.35$, the He burning luminosity is $L_{\text{He}} \approx 0.18L_{\text{H}}$, and the accretion rate is given by

$$\dot{m}_{\text{acc}} \approx 3.45 \times 10^{-8} \left(\frac{L_{\text{nuc}}}{10^{37} \text{ erg s}^{-1}} \right) \text{ M}_{\odot} \text{ yr}^{-1} \quad (2.75)$$

(see also Iben, 1977). Because L_{nuc} is much larger than the accretion luminosity L_{acc} (see §2.3.3), the total luminosity can be approximated as being equal to L_{nuc} , although it also contains a relatively small contribution from L_{acc} .

When the accretion rate $\dot{m}_{\text{acc}}$ is smaller than $\dot{m}_{\text{std}}$, steady burning is not possible and hydrogen burning will take place as unstable flashes or novae, as in the classical novae (CN) and recurrent novae (RN). Loci of constant ignition mass $m_{\text{env}}^{\text{crit}}$ are indicated in Fig. 2.12 for $\dot{m}_{\text{acc}} < \dot{m}_{\text{std}}$. These shell flashes can last from ~ 1 month to years (Prialnik and Kovetz, 1995; Kato, 1997, 1999). A fraction of the envelope mass will be ejected during the nova outburst. Lower accretion rates are associated with more violent outbursts, and the ejection of more material, and it is expected that for $\dot{m}_{\text{acc}} < 0.25\dot{m}_{\text{std}}$, all accreted matter will be ejected (Kahabka and van den Heuvel, 2006). However, the amount of material that is ejected in reality, and the implications for the change in the WD mass is still controversial (e.g. Starrfield, 2015).

After the start of a nova explosion, the WD envelope expands considerably and mass loss occurs through an optically thick wind (e.g. Kato, 2010). As a large part of the envelope is blown off, the photospheric radius shrinks and the effective temperature rises. The companion will re-appear from the shrinking WD envelope and an accretion disc may be formed again. After the optically thick wind has ceased, supersoft X-rays will be emitted from the system up to the point when the hydrogen burning stops.

Many novae have indeed been observed as supersoft sources during this stage of their outbursts, with an observed supersoft X-ray phase of $\lesssim 10$ yr (Orio, 2004). A well-known example is the classical nova GQ Mus (Ögelman et al., 1993). Nova V1974 Cygni appeared as an SSS until ~ 18 months after the outburst commenced, when the accreted hydrogen fuel was exhausted (Krautter et al., 1996). Nova LMC 1995 appeared as an SSS from 5 months up to 3 years after its outburst (Orio and Greiner, 1999).

It had been a long-accepted fact in binary evolution models that the hydrogen-rich envelope on a WD surface will expand to the size of a red giant when the accretion rate exceeds the critical limit $\dot{m}_{\text{cr}}$ (e.g. Nomoto, Nariai and Sugimoto, 1979; Nomoto,

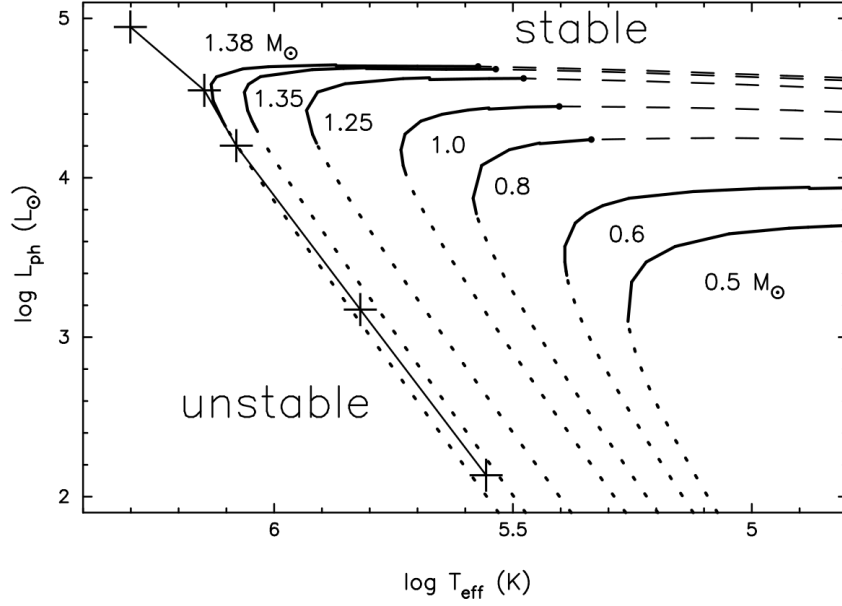


Fig. 2.13: An HR diagram (luminosity versus effective temperature) for accreting WDs, with loci of constant WD mass. The dotted lines indicate regions associated with unstable flashes ($\dot{m}_{\text{acc}} < \dot{m}_{\text{std}}$), while the dashed and solid lines indicate regions where stable burning can take place ($\dot{m}_{\text{acc}} > \dot{m}_{\text{std}}$). The dashed lines represent regions where an optically thick accretion wind will blow ($\dot{m}_{\text{acc}} > \dot{m}_{\text{cr}}$). (Adopted from Kato, 2010, Fig. 1)

1982). As a result of viscous drag, the two stars in this common envelope will spiral in towards each other, and a double-degenerate system is formed (e.g. Iben and Tutukov, 1984; Webbink, 1984).

However, Hachisu, Kato and Nomoto (1996) found that a strong so-called “accretion wind” ($\sim 1000 \text{ km s}^{-1}$) starts to blow from the WD surface once $\dot{m}_{\text{acc}} > \dot{m}_{\text{cr}}$. This prevents the formation of a common envelope and therefore keeps the binary from collapsing. Steady hydrogen burning can still take place, and the helium layer on its surface can grow at a rate of $\dot{m}_{\text{He}} \approx \dot{m}_{\text{acc}}$. The mass loss rate through the wind is then $\dot{m}_{\text{wind}} \approx \dot{m}_{\text{acc}} - \dot{m}_{\text{cr}}$, and may typically be $\lesssim 10^{-6} M_{\odot} \text{ yr}^{-1}$. When this wind is optically thick, the observation of supersoft X-rays is not expected, due to self-absorption. In Fig. 2.13, the different regimes associated with accreting, hydrogen burning WDs are shown on an HR diagram.

The luminosity and mass of a hot, degenerate C-O WD that is powered by hydrogen shell burning is estimated by

$$L = 2.3 \times 10^{38} \left(\frac{M_{\text{wd}}}{M_{\odot}} - 0.52 \right) \text{ erg s}^{-1} \quad (2.76)$$

and plotted as the dash-dotted curve in Fig. 2.14 (see Iben and Tutukov, 1996 and references therein). This luminosity represents the “plateau” luminosity at the

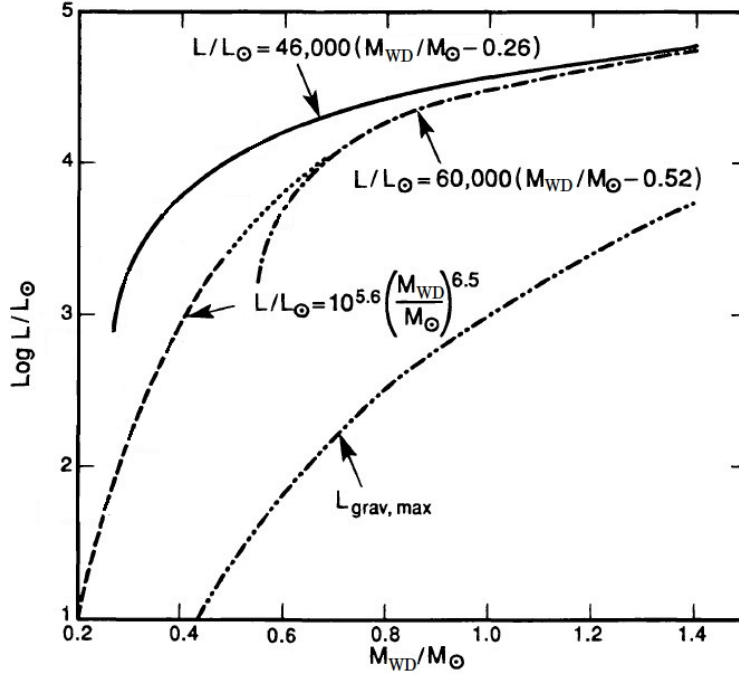


Fig. 2.14: The mass-luminosity relation for degenerate WDs with hydrogen shell burning. The solid curve is the plateau luminosity curve for cool WDs. The dashed curve is for He WDs, and the dash-dotted curve for hot C-O WDs. The dash-double-dotted curve illustrates $L_{\text{grav,max}}$, the maximum luminosity that can be derived from accretion only by WDs with different masses. (Adopted from Iben and Tutukov, 1996, Fig. 5.)

horizontal track in Fig. 2.13. For a He WD, the relation is

$$L = 1.5 \times 10^{39} \left(\frac{M_{\text{wd}}}{M_{\odot}} \right)^{6.5} \text{ erg s}^{-1} \quad (2.77)$$

during the plateau hydrogen burning phase after its departure from the first giant branch, as shown by the dashed curve in Fig. 2.14. When a zero-temperature WD consisting of He, C and O (or O and Ne) is considered, its luminosity during the hydrogen burning following a shell flash can be obtained with

$$L = 1.8 \times 10^{38} \left(\frac{M_{\text{wd}}}{M_{\odot}} - 0.26 \right) \text{ erg s}^{-1} . \quad (2.78)$$

According to Iben and Tutukov (1996), SSSs that have been undergoing stable burning for some time normally have hot cores and are described quite well by Eq. (2.76). WDs in CVs and symbiotic sources undergoing nova outbursts are more likely to be old, and the approximation for a cool WD in Eq. (2.78) should be used.

2.6.4 Long-term variability

CAL 83 in the LMC was for long considered to be the prototype of steady burning SSSs, i.e. with $\dot{m}_{\text{std}} < \dot{m}_{\text{acc}} < \dot{m}_{\text{cr}}$ as discussed in the previous section. However, this source and also several other SSSs have been shown to exhibit significant X-ray

variability on superorbital time-scales (see Chapter 3 for a discussion of CAL 83 and the similar LMC source RX J0513.9-6951). The time-scales of these variations have been used by various authors to constrain the source properties.

For accretion rates above the limit of $\dot{m}_{\text{cr}}$, Van den Heuvel et al. (1992) proposed a “limit cycle” model. The excess accumulated mass piles up, increasing the WD photospheric radius and significantly decreasing the effective temperature, while hydrogen burning still takes place on the surface. This causes the formation of a “red giant envelope”, with a mass of $\sim 10^{-5} M_{\odot}$. The giant configuration fills the primary Roche lobe and its outer parts surround the secondary star. Accretion will be suppressed and material transferred from the secondary probably ejected via the L₂ point.

Within a time-scale of ~ 100 years, the nuclear fuel on the WD surface will be exhausted before spiral-in can occur, the envelope collapses and accretion is re-established, restarting the cycle. It is estimated that the source undergoing such a limit cycle will be in the “on”-state in soft X-rays $\gtrsim 10\%$ of the time. However, in view of the existence of an accretion wind preventing the formation of the “red giant” configuration (Hachisu et al., 1996), the Van den Heuvel et al. (1992) limit cycle may not be able to develop.

Kahabka (1995) investigated the X-ray variability in the transient/recurrent SSSs RX J0513.9-6951, RX J0527.8-6954, RX J0019.8+2156 and RX J0045.4-4154, and considered the possibility of the X-ray “outbursts” being caused by recurrent nuclear burning shell flashes. This would place these sources in the accretion rate regime slightly below the steady burning limit ($\dot{m}_{\text{std}}$), but above the accretion rate leading to nova explosions, i.e. in a transition region between recurrent novae and steady SSSs. Such “recurrent SSSs” would not exhibit the large-scale mass ejection associated with nova outbursts, although mass loss due to winds may well occur.

The measurable time-scales in a recurrent SSS are the X-ray decay time t_{decay} , and the recurrence time t_{recur} of the cycle. Therefore, an accreting WD would accrete material until the critical envelope mass $m_{\text{env}}^{\text{crit}}$ is reached, nuclear burning will commence and continue for a time t_{decay} until the envelope has been completely consumed by nuclear burning and possibly also mass loss through winds. Consequently, accretion without nuclear burning will take place for a time $t_{\text{recur}} - t_{\text{decay}}$ before the critical envelope mass is reached again.

Consider a WD accreting at a rate $\dot{m}_{\text{acc}}$, with $\dot{m}_{\text{stable}}$ the rate of nuclear burning on its surface and $\dot{m}_{\text{wind}}$ the constant rate of mass loss through a wind. With a straightforward analytical approach, Kahabka (1995) showed that the critical envelope mass

can be expressed as

$$m_{\text{env}}^{\text{crit}} = \frac{\dot{m}_{\text{stable}} t_{\text{decay}} \left(1 - \frac{t_{\text{decay}}}{t_{\text{recur}}}\right)}{1 - f_e \left(1 - \frac{t_{\text{decay}}}{t_{\text{recur}}}\right)} \quad (2.79)$$

where f_e represents the fraction of the envelope that is ejected during a shell flash. For shell flashes in recurrent SSSs, f_e is expected to be near 0, while during nova outbursts, f_e should be close to 1. Interestingly, Eq. (2.79) is independent of $\dot{m}_{\text{wind}}$, which just acts to reduce the effective accretion rate. The value of $\dot{m}_{\text{stable}}$ can be estimated by $\dot{m}_{\text{std}} \lesssim \dot{m}_{\text{stable}} \lesssim \dot{m}_{\text{cr}}$, and after obtaining $m_{\text{env}}^{\text{crit}}$ from the observable time-scales, an estimate of the WD mass can be determined with Eq. (2.71).

The actual turn-on in soft X-rays have also been modelled as the extended WD envelope contracts after the initiation of a nuclear flash (Kahabka, 1996c and references therein), resulting in an effective temperature increase from ~ 100000 to ~ 500000 K. The length of the observed turn-on phase in X-rays has been expressed analytically as

$$t_{\text{contr}}^{\text{obs}} \approx 70 \left[\frac{m_{\text{env}} (1 - f_{\text{loss}})}{4 \times 10^{-6} M_{\odot}} \right] \left(\frac{R_{\text{wd}}}{5 \times 10^8 \text{ cm}} \right)^{-1} \text{ d} , \quad (2.80)$$

with f_{loss} the fraction of the accreted envelope that is ejected either during the outburst or in a wind during the contraction phase. The envelope mass m_{env} can be approximated by the critical envelope mass $m_{\text{env}}^{\text{crit}}$.

Another related time-scale is the Kelvin-Helmholtz time-scale of the envelope, i.e.

$$\tau_{\text{KH}} \approx 1200 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right) \left(\frac{m_{\text{env}}}{4 \times 10^{-6} M_{\odot}} \right) \left(\frac{R}{10^9 \text{ cm}} \right)^{-1} \left(\frac{L}{10^{37} \text{ erg s}^{-1}} \right)^{-1} \text{ d} , \quad (2.81)$$

which determines the time-scale for the contraction of the envelope after an outburst (see Alcock et al., 1997; Kahabka and van den Heuvel, 2006 and references therein).

2.6.5 The X-ray spectrum

Blackbody models

When the gas in a source of radiation is in thermodynamic equilibrium, its radiated spectrum depends only on the temperature, and the specific intensity of radiation at frequency ν is given by (e.g. Irwin, 2007, p. 123)

$$I_{\nu} = B_{\nu}(T) = \frac{2h\nu^3}{c^2} \frac{1}{e^{h\nu/kT} - 1} . \quad (2.82)$$

This is known as *blackbody radiation*. The bolometric surface flux of a blackbody emitter is related to the temperature by the Stefan-Boltzmann law, i.e.

$$F_{\text{bol}} = \sigma T^4 , \quad (2.83)$$

but can also be expressed as

$$F_{\text{bol}} = \frac{L_{\text{bol}}}{4\pi R^2} , \quad (2.84)$$

where L_{bol} is the bolometric luminosity of the source, and R its radius, therefore

$$L_{\text{bol}} = 4\pi R^2 \sigma T^4 . \quad (2.85)$$

Blackbody models were initially used to fit the low-resolution X-ray spectra of SSSs obtained with *EXOSAT*, *ROSAT*, *BeppoSAX* and *ASCA*. Statistically, these blackbody fits were quite satisfactory, but the derived bolometric luminosities were often larger than the Eddington limit for a WD. Also, after the first high-resolution (~ 0.05 Å) spectra of SSSs were obtained with *XMM-Newton* and *Chandra*, it was also obvious that the fine structure of the spectrum, including various absorption features and also emission lines, can not be fit by a smooth continuum spectrum to a high degree of accuracy (e.g. Paerels et al., 2001; Lanz et al., 2005; Ness et al., 2013). This casts doubt whether pure blackbody models are adequate to model the X-ray emission from SSSs.

White dwarf atmosphere models

After the accreting WD model for SSSs was proposed, white dwarf atmospheres were explored to model the X-ray data. Nuclear burning occurs at a large optical depth below the photosphere of the WD, with temperatures $\sim 10^8$ K. However, photo-absorption by the highly ionized metals in the WD atmosphere causes absorption edges in the spectrum, and the X-ray spectrum of a WD deviates significantly from a blackbody (e.g. Heise, van Teeseling and Kahabka, 1994). In particular, a WD atmosphere model emits a larger fraction of its bolometric luminosity in the super-soft X-ray band than a blackbody spectrum of the same temperature. Although a blackbody model and a WD atmosphere model may yield a similar goodness of fit in the supersoft X-ray band when analysing low-resolution data, integrating the fitted models over all frequencies will then yield a bolometric luminosity for the WD atmosphere model that is approximately an order of magnitude lower than that of the blackbody, and somewhat lower than the Eddington limit. WD atmosphere model fits also yield slightly smaller, more realistic values for the radius of the hydrogen burning WD.

Heise et al. (1994) considered models in local thermodynamic equilibrium (LTE), assuming that the high density associated with the high WD surface gravity g causes the atmosphere to be in thermodynamic equilibrium. The LTE models have a much flatter shape than a blackbody, with a cut-off at high energies because of several absorption edges. However, the LTE assumption might not be valid in the outer, low-density region of the atmosphere, where departures from LTE also need to be considered. Models incorporating the latter are called non-LTE (NLTE) models. Hartmann and Heise (1997) showed that NLTE models differ significantly from LTE models for $\log g < 9$. Also, NLTE models are not cut off as strongly as the LTE models at high energies, because the depths of the edges in NLTE are decreased, or even in emission.

The vast increase in available computational power during the last few decades have enabled major increases in the sophistication of the computed WD atmosphere models (see Rauch and Werner, 2010 and references therein). These models involve the solution of the equations of radiative and hydrostatic equilibrium. In addition, the current modelling codes enable the generation of model spectra with not only the bound-free element opacities, but also the effects of line opacities and “full metal-line blanketing”, enabling even more accurate fits, and also the identification of spectral lines in the observed spectra. For example, the Tübingen NLTE Model-Atmosphere Package (TMAP) is a simulation code for the calculation of NLTE stellar atmospheres in spherical or plane-parallel geometry in hydrostatic and radiative equilibrium, including all elements from hydrogen to nickel⁴ (Werner and Dreizler, 1999; Werner et al., 2003; Rauch and Deetjen, 2003). It also has a user-friendly interface where specified model grids can be requested even by inexperienced users⁵.

In order to fit WD atmosphere spectra to an observed X-ray spectrum, one first needs to obtain a grid of simulated spectra, usually for several values of the surface gravity $\log g$ and effective temperature T_{eff} . (The surface gravity of WDs range from $\sim 10^7$ for lower mass WDs, to an upper limit of $10^{10} \text{ cm s}^{-2}$ for the massive O-Ne-Mg WDs, e.g. Hartmann and Heise, 1997.) For a fixed set of element abundances, ionization states, energy levels and lines included, the shape of the unabsorbed simulated spectrum is determined by $\log g$ and T_{eff} . A multiplicative factor in the model should also provide for the interstellar absorption, quantified by the hydrogen column density N_{H} . In addition, a normalization parameter $(R/d)^2$ is obtained from the fit, where R is the source radius, and d the distance to the source.

After obtaining $\log g$, T_{eff} and R (if d is known) from the fit, the mass of the WD

⁴<http://www.uni-tuebingen.de/de/41621>

⁵<http://astro.uni-tuebingen.de/~TMAW>

can be obtained with

$$M_1 = \frac{10^{\log g} R^2}{G} \quad (2.86)$$

and the luminosity from Eq. (2.85). It is important to bear in mind that the effective X-ray photosphere will be somewhat larger than that of the degenerate core, therefore the direct substitution of R into Eq. (2.25) will cause an underestimation of the WD mass. In fact, the photospheric radius in SSSs can be ~ 2 -3 times larger than the WD core (Ibragimov et al., 2003 and references therein).

Ibragimov et al. (2003) fitted line-blanketed LTE models to the ROSAT spectra of 11 SSSs. Suleimanov and Ibragimov (2003) performed a detailed investigation of the derived parameters of 10 of these sources in the context of the stable surface burning model, and inferred the following relation between the mass and effective temperature of SSSs in the stable burning strip:

$$T_{\text{eff}} = 3 \times 10^5 \text{ K} \frac{\frac{M_{\text{wd}}}{M_{\odot}} \left(\frac{M_{\text{wd}}}{M_{\odot}} - 0.26 \right)^{0.25}}{\left[1 - 1.53 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right) + 1.33 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right)^2 - 0.51 \left(\frac{M_{\text{wd}}}{M_{\odot}} \right)^3 \right]^{0.5}}. \quad (2.87)$$

It is assumed that $R = 2(M_{\text{wd}}/M_{\odot})^2 R_{\text{wd}}$, and that the luminosity is equal to one half of the plateau luminosity.

The interaction of soft X-rays with the interstellar medium

A very important aspect in the study of SSSs is the fact that the interstellar medium (ISM) is only partially transparent at soft X-ray energies. The detectability of SSSs and the shape of the absorbed spectrum is therefore strongly dependent on the column density in the direction of the source. The photoelectric absorption of X-rays in the ISM is characterized by sharp increases at the K and L edges of the elements present in the ISM, and the overall absorption is approximately proportional to $E^{-8/3}$ (e.g. Gorenstein and Tucker, 1976 and references therein).

For a hydrogen column density of $N_{\text{H}} = 10^{20} \text{ cm}^{-2}$, the amount of absorption above 0.4 keV is negligible. However, the attenuation of the spectrum at 0.2 keV is approximately one optical depth and there is only a tiny amount of transmission below 0.1 keV. For $N_{\text{H}} = 5 \times 10^{21} \text{ cm}^{-2}$, almost no radiation is transmitted below 0.3 keV, and when $N_{\text{H}} > 10^{22} \text{ cm}^{-2}$, there is essentially no transmission below 1.5 keV, and SSSs are not detectable. This explains why the detection rate of SSSs in the Galactic plane is very low. However, it should be noted that this simple picture of absorption can be significantly altered by e.g. the presence of a substantial amount of H_2 , ionized H and He or dust grains along the line of sight (see also Overbeck, 1965; Field, 1974).

An accurate estimate of the amount of neutral hydrogen along the line of sight is absolutely essential to allow accurate modelling of the low energy (<1 keV) X-ray properties of these systems.

Multiwavelength modelling

Skopal (2013, 2015a,b,c) introduced a method involving the multiwavelength modelling of the spectral energy distribution of SSSs, i.e. fitting the whole spectrum from the near-infrared to the X-ray band simultaneously. The full details of these models will not be discussed here, but it is noted that they have the major advantage that they overcome the problem of the mutual dependence between T_{eff} , L_{bol} and N_{H} that often complicate fitting conclusive models to X-ray spectra alone. A multi-wavelength approach allows an unambiguous solution, and can therefore provide a much more accurate description of the object.

2.7 Instabilities and turbulence in accretion discs

While rotation and pulsation of the accreting WD constitute diagnostic tools to study the accretion process and response of the WD to accretion, other temporal observational signatures may also be introduced in the accretion disc as a result of turbulence. For the observational astronomer, these signatures may allow decoding the complex plethora of processes that drive the high accretion rate in SSSs. Therefore, a slightly more quantitative discussion will be presented here.

As explained in §2.4.2, some form of turbulence driven by viscosity is required in an accretion disc to provide a rate of outward angular momentum transport that is sufficient to explain the accretion rates inferred from observations. The origin of this turbulence has been a mystery for a long time. Hydrodynamic turbulence in the form of the classical shear-flow instability or convective turbulence seemed to be the most obvious potential causes, but it turns out that these are very ineffective in an accretion disc. However, Balbus and Hawley (1991) developed a theory of turbulence driven by magnetohydrodynamic (MHD) instabilities, which provides an effective mechanism of angular momentum transport in a disc.

This section starts off with a summary of the typical time-scales of variation in the disc (based on Frank et al., 2002, pp. 110-113), followed by an explanation of hydrodynamic and also magnetohydrodynamic turbulence (mostly based on the discussions of Biskamp, 2003, pp. 233-255 and the review of Balbus and Hawley, 1998), and concludes with a discussion of observed variability that may be related to accretion disc turbulence.

2.7.1 Disc time-scales

The shortest characteristic time-scale of the disc is the *dynamical time-scale*

$$t_\phi \sim \frac{R}{v_\phi} \sim \Omega_K^{-1}(R) . \quad (2.88)$$

The time-scale on which material diffuses through the accretion disc under the influence of viscous torques is the *viscous time-scale*

$$t_{\text{visc}} \sim \frac{R^2}{\nu} \sim \frac{R}{v_R} , \quad (2.89)$$

while the *thermal time-scale* describes the time required for a re-adjustment to thermal equilibrium:

$$t_{\text{th}} = \frac{\text{heat content per unit disc area}}{\text{dissipation rate per unit disc area}} . \quad (2.90)$$

It can be proven that the thermal time-scale may also be expressed as

$$t_{\text{th}} = \left(\frac{c_s}{v_\phi} \right)^2 t_{\text{visc}} . \quad (2.91)$$

It can be shown that, for the α -disc solutions, the dynamical and viscous time-scales are

$$t_\phi \sim \alpha t_{\text{th}} \sim 100 \left(\frac{M_1}{M_\odot} \right)^{-1/2} \left(\frac{R}{10^{10} \text{ cm}} \right)^{3/2} \text{ s} \quad (2.92)$$

and (2.93)

$$t_{\text{visc}} \sim 4 \times 10^4 \alpha^{-4/5} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} M_\odot \text{ yr}^{-1}} \right)^{-3/10} \left(\frac{M_1}{M_\odot} \right)^{1/4} \left(\frac{R}{10^{10} \text{ cm}} \right)^{5/4} \text{ s} . \quad (2.94)$$

respectively. These time-scales govern most of the dynamic processes of the disc driven by MHD turbulence.

2.7.2 Hydrodynamic turbulence

The hydrodynamic shear-flow stability of an incompressible, rotating fluid is governed by Rayleigh's criterion, i.e.

$$\frac{d(R^2\Omega)^2}{dR} > 0 . \quad (2.95)$$

From Eq. (2.34), it is easy to see that Rayleigh's criterion is satisfied in a Keplerian disc. Stabilizing Coriolis forces will therefore smooth out shear-flow instabilities

on the dynamical time-scale of the disc. Usually, “instability” refers to a linear instability, in other words the growth of an infinitesimal perturbation to a larger scale.

Non-linear instabilities, on the other hand, are caused by perturbations that are finite, even though still small. It is known that non-rotating flows are in many cases nonlinearly unstable, for example the Couette flow in a fluid between two parallel plates with relative movement, and the Poiseuille flow in a fluid flowing through a pipe. Even though these are linearly stable, they become fully turbulent once the perturbation exceeds a certain threshold (e.g. Orszag and Kells, 1980). This suggests that non-linear hydrodynamic instabilities may be the cause of accretion disc turbulence. However, simulations have shown that this is not the case, and thus that Keplerian accretion discs are stable to both linear and non-linear hydrodynamic instabilities (Balbus, Hawley and Stone, 1996).

The accretion disc will have a temperature gradient in the z -direction, decreasing from the central temperature towards the lower temperature at the surface of the disc. Convective turbulence is expected to occur because of this “vertical” temperature gradient. Because the accretion process leads to the conversion of gravitational potential energy to heat, it would lead to an increase in the midplane temperature, yielding steeper temperature gradients and enhanced convective turbulence. Therefore, *if* convective turbulence facilitated the outward transfer of angular momentum, it could yield a sustainable process. However, numerical simulations show that the transport of angular momentum by convective turbulence in an accretion disc is extremely weak, and also that it causes an *inward* transfer of angular momentum, essentially disqualifying it as the driving agent of accretion (Stone and Balbus, 1996).

2.7.3 Magnetohydrodynamic turbulence

The accretion discs in compact binaries and also those found in active galactic nuclei (AGN), are hot enough to be significantly or fully ionized. Considering the close connection between electricity and magnetism, the free charges in the disc implies a strong coupling between the disc material and any magnetic fields that may be present. Charged particles can readily flow along magnetic field lines, but can not easily cross the field lines. Magnetic field lines are also dragged along by the motion of ionized material. Therefore, the magnetic field lines and the ionized material are “frozen” together, and tend to move together. As a result, a Keplerian disc is unstable to magneto-hydrodynamic turbulence, as shown by Balbus and Hawley (1991).

Consider two fluid elements at different radii in the disc that are connected by a magnetic field line. The element at smaller R has a higher angular velocity than the outer element, causing the field line to become “stretched”, analogous to a weak

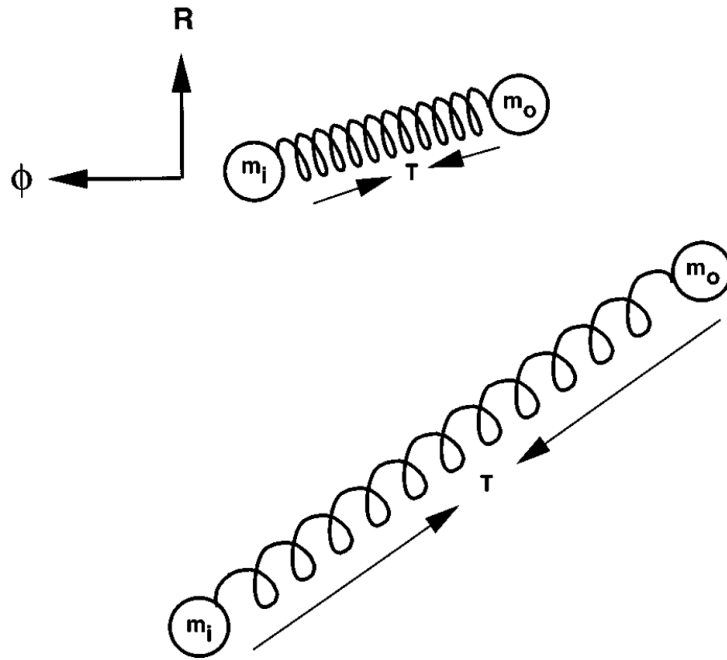


Fig. 2.15: The weak spring analogy for a magnetic field line. Two masses are shown orbiting in the plane of the accretion disc: the inner mass m_i and the outer mass m_o . The tension force T exerted by the spring results in the transfer of angular momentum from m_i to m_o . The result is that m_i drops into a faster orbit closer to the primary, and moves further ahead, while m_o moves into a slower orbit farther away, and falls further behind. T increases, and the process runs away. (Adopted from Balbus and Hawley, 1998, Fig. 16.)

spring that is stretched (see Fig. 2.15). This increased tension in the field line causes a braking action on the inner element, and an acceleration of the outer element, effectively transferring momentum from the inner to the outer element. However, because angular momentum increases outwards in a Keplerian disc, the resultant effect is that the inner element moves into an orbit even closer to the primary, with a higher angular velocity, and the outer element moves into an even wider orbit with a lower angular velocity, and this stretches the field line even more. The stretching of the field line is equivalent to strengthening the field, so the initial fields are amplified. The mechanism described here is known as the *magnetorotational instability* (MRI), or also the *Balbus-Hawley instability* (Balbus and Hawley, 1991, 1998).

As the field line is increasingly stretched, reconnection will ultimately take place to dissipate energy, resulting in “bubbles” of gas being transported to different radii, and the ordered Keplerian flow breaks up into turbulence (Fig. 2.16). This instability develops rapidly and is very powerful, although the magnetic field should not be too strong, i.e., the magnetic field energy density must be less than the thermal energy density.

The magnetic field causing the onset of the MRI may be the magnetic field of the

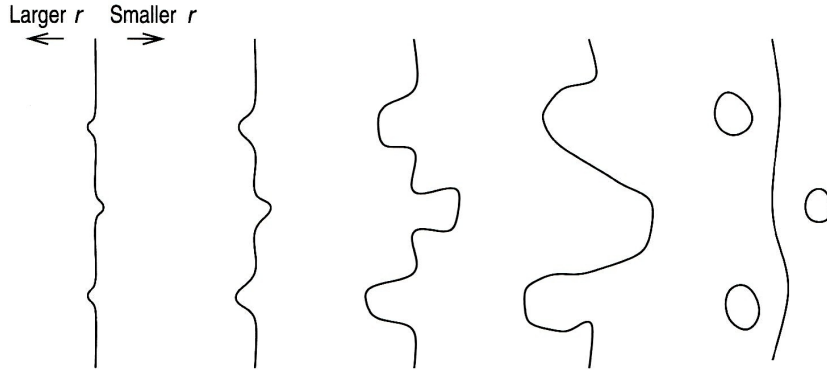


Fig. 2.16: The sequence from the left hand side to the right hand side shows the growth of a Balbus-Hawley instability in a magnetic field line perpendicular to the disc plane. The initial small deviations from a straight line are amplified, and the field strength is increased until reconnection occurs. (Adopted from Hellier, 2001, Fig. 5.6, p. 63.)

primary itself, or possibly fields advected from the secondary star. However, it has also been shown that the shear-flow turbulence in the disc amplifies the existing fields by the dynamo effect (e.g. Campbell, 1997, pp. 211-248). The process of magnetic field amplification can reach saturation due to the vertical convection of the field into the accretion disc corona, or by reconnection and field dissipation, with the latter processes being the dominant ones.

In turbulent flows, it is observed that large-scale eddies subdivide into smaller and smaller eddies. This process during which the kinetic energy of large-scale eddies is successively subdivided until it contributes to the thermal motion of particles is known as a Kolmogorov cascade (e.g. Choudhuri, 1998, pp. 158-175). The associated increase in the disc temperature and therefore the viscosity also enhances the outward angular momentum transfer.

The Kolmogorov process typically takes place on time-scales between ~ 100 and ~ 1000 s. It has been explained that hydrodynamic turbulence is suppressed on the disc dynamical time-scale. However, if the Kolmogorov cascade can thermalize the turbulent kinetic energy before the turbulence is smoothed out, hydrodynamic turbulence can still boost the disc temperature. From Eq. (2.92), one can see that for a solar-mass WD, the dynamical disc time-scale is of the order of 100 and 1000 s at radii of 10^{10} cm and 5×10^{10} cm respectively.

Therefore, the dynamical time-scale can be longer than the time required for a complete Kolmogorov cascade in the outer regions of the disc. This is especially relevant because of the existence of the bright spot, where the collision of the accretion stream with the disc is expected to create considerable hydrodynamic turbulence.

In the next section, the nature of flickering and quasi-coherent oscillations in WD binaries will be discussed. Although these transient phenomena are usually observed in cataclysmic variables (CVs), they may also be observable in SSSs, even though the latter have much higher accretion rates.

2.7.4 Flickering and quasi-coherent oscillations

The lightcurves of many CVs have been observed to exhibit random variability on time-scales of seconds to hours, and this is often referred to as *flickering* (e.g. Hellier, 2001, pp. 151-153; Scaringi, 2014). From studies of certain eclipsing CVs, it seems that the flickering mostly originates in the inner accretion disc, or sometimes at the bright spot. The longer-lasting variations have larger amplitudes than the shorter ones, making disc turbulence a very probable cause of flickering. Larger turbulent eddies will be brighter, and their “swirling motion” will also take longer, i.e. their turnover frequency is smaller.

Consequently, a Fourier spectrum of a flickering source should exhibit a “red noise” signature, with strong power at lower frequencies associated with large eddies, decreasing to weaker power values at higher frequencies associated with smaller eddies. In fact, this red noise phenomenon can be quantified quite elegantly. Each turbulent cell has an associated wavenumber k that is inversely proportional to the size of the cell. During the Kolmogorov cascade, the turbulent energy is dependent on k through $E(k) \propto k^{-5/3}$, which is sometimes referred to as the Kolmogorov law. As shown by Biskamp (2003, p. 102), the Kolmogorov law not only applies to hydrodynamic turbulence, but also to magnetohydrodynamic turbulence. To be exact, $E(k) \propto k^{-5/3}$ in a direction perpendicular to the magnetic field, while $E(k) \propto k^{-5/2}$ parallel to the field. The red noise regime of the power spectrum can thus be expected to have a slope of approximately $-5/3$. Therefore a fit through the red noise part of a Fourier power spectrum should reveal this $k^{-5/3}$ relationship if hydrodynamic or MHD turbulence is present.

In addition to random flickering, many CVs also exhibit variability with a preferred time-scale, but which is not strictly periodic. These oscillations have periods between a few seconds and a few thousand seconds (see e.g. the summary of Hellier, 2001, pp. 153-157, and the review papers of Warner, 2004 and Warner and Woudt, 2008). The first type of these oscillations were discovered in dwarf novae during outburst, and thus became known as dwarf nova oscillations (DNOs).

The quiescent state of dwarf novae is associated with a cold disc that is only partially ionized, in which the viscosity and associated accretion rate through the disc is low (e.g. Hellier, 2001, pp. 62-64). If the mass transfer rate from the donor is

greater than the rate at which material can flow through this low-viscosity disc, the surface density and associated temperature of the disc will increase. At a certain critical temperature, rapid ionization occurs and the disc temperature and viscosity rises sharply, causing a dramatically enhanced accretion rate and luminosity. The accretion through the disc now occurs at a higher rate than the mass transfer rate from the secondary, and the material in the disc is drained onto the WD, causing the disc surface density and temperature to decrease. At another critical point, rapid de-ionization will occur and the disc returns to its cold, quiescent state.

Apart from their occurrence in dwarf novae during outburst, DNOs are also found in other CVs with high accretion rates, but they are noticeably absent in almost all intermediate polars. (Note that, in this context, a “high” accretion rate does not necessarily refer to the very high $\dot{m}_{\text{acc}}$ needed for steady nuclear burning as discussed in §2.6.3, but just to accretion rates higher than that of dwarf novae during quiescence, i.e. as found in dwarf novae during outburst and in novalikes.) They have periods in the $\sim 5\text{--}40$ s range, usually with a variation amplitude of $< 1\%$ in the optical. The oscillation amplitude can be variable on time-scales of hours, and can also disappear and reappear. They have also been observed in soft X-rays and in the extreme ultraviolet, with a much larger amplitude, even up to 100%.

DNOs exhibit a period-luminosity relationship, in the sense that the minimum period corresponds to the maximum accretion rate during the dwarf nova outburst, and the period increases during the after-outburst decline. They sometimes exhibit small jumps in the period itself ($\sim 0.01\%$). Two DNO periods are sometimes observed simultaneously.

The rate of change of a period can be characterized by the *quality factor*, given by

$$Q = \left| \frac{dP}{dt} \right|^{-1}, \quad (2.96)$$

i.e. fast changes in the period yields a small value for Q . For example, for the coherent spin period detection in the intermediate polar DQ Her, $Q \sim 10^{12}$ (see §2.6.1). For DNOs, Q is usually between 10^3 and 10^7 .

The second type of oscillation exhibits similar characteristics to DNOs, but, with periods typically ~ 4 times longer than those of DNOs, these oscillations are called “longer period DNOs” (lpDNOs) (Warner et al., 2003). They are also observed only in CVs with high accretion rates, but without the strong period-luminosity relationship of DNOs, and with slightly larger amplitudes. In cases where DNOs and lpDNOs are observed simultaneously, their amplitude and phase modulations are independent.

The third type is referred to as “quasi-periodic oscillations” (QPOs). Their periods and modulation amplitudes are much larger than those of DNOs, and they are less coherent. Typically, $5 < Q < 20$, and harmonics of the fundamental frequency are often seen as well. QPOs have been observed in CVs with high accretion rates (i.e. the accretion rates typical of novalikes and dwarf novae in outburst), but also in dwarf novae during quiescence, and they are not dependent on the simultaneous presence of DNOs.

However, in cases where QPOs do occur simultaneously with DNOs, it is found that $P_{\text{QPO}}/P_{\text{DNO}} \sim 15$, and the QPO is then referred to as a “DNO-related QPO” (Warner et al., 2003). In several systems with “double DNOs”, the QPO is the beat period⁶ of the double DNOs.

A second type of QPO also exists, having much longer periods of $\sim 1000\text{--}3000$ s. Many CVs exhibiting a QPO with $P \sim 1000$ s are suspected to be intermediate polars, in which the coherent WD rotation is not observed directly, possibly due to a high accretion rate, but reprocessed by a QPO source with varying period. This is supported by these QPOs having periods comparable to that of the WD rotation periods in canonical intermediate polars, i.e. ~ 15 min. However, in cases where QPOs are observed simultaneously with DNOs, the system can not be an intermediate polar (see below). In these systems, the longer period QPOs may represent the Keplerian periods of blobs at the outer edge of an accretion disc. Another source of QPOs may be a modulation in mass transfer caused by non-radial oscillations of the secondary.

What causes the rich variety of DNOs in CVs? The most favoured interpretation is that they may originate from a region close to the surface of a WD with a magnetic field too small to enable rigid body rotation by the core and exterior region, but with a lower limit still high enough to control accretion close to its surface.

Katz (1975) studied the rotation period of the WD in DQ Her, and argued that the clock-like stability in the WD rotation in intermediate polars must be due to rigid-body rotation of the WD. He proposed the WD magnetic field as the mechanism coupling the stable clock of the core to the exterior regions. The magnetic field that is necessary to transmit an observed acceleration $\dot{\Omega}_*$ from the outer regions to the core, or the other way around, is given by comparing the Maxwell and mechanical

⁶If a period P_{beat} is the beat period generated by two other periods P_1 and P_2 , then $\frac{1}{P_{\text{beat}}} = \left| \frac{1}{P_1} - \frac{1}{P_2} \right|$

stresses:

$$\frac{B_r B_\phi}{4\pi} \approx \rho_{\text{wd}} \dot{\Omega}_* R^2, \quad (2.97)$$

where B_r and B_ϕ are the radial and azimuthal components of the internal magnetic field. By estimating $R \sim 10^9$ cm and a density of $\rho_{\text{wd}} \sim 10^6$ g cm $^{-3}$ (assuming that a substantial portion of the deep interior is still in fluid form and not yet crystalline), together with $\dot{\Omega}_* \approx 10^{-15}$ s $^{-2}$ measured for DQ Her, $B_r B_\phi \sim 10^{10}$ G 2 was obtained, yielding a minimum field strength of $B_r = B_\phi \sim 10^5$ G.

A magnetic field weaker than this will not be able to couple the accretion torque to the interior, and this will allow the accumulation of an equatorial belt rotating rapidly at the local Keplerian velocity, i.e. faster than the WD core. Instead of the period of the equatorial belt being exactly equal to the local Keplerian period, its inertia may cause it to have a somewhat longer period than this (Warner, 1995*b*). Evidence for the existence of such an equatorial belt has been found in UV spectra of the dwarf novae VW Hyi, U Gem, AL Com and RU Peg (Sion et al., 1996; Gänsicke and Beuermann, 1996; Cheng et al., 1997; Szkody et al., 1998; Sion and Urban, 2002). This concept was initially suggested by Paczyński (1978), and was developed by Warner (1995*b*) and Warner and Woudt (2002) into a low-inertia magnetic accretor (LIMA) model to explain the DNO phenomenon, which will be briefly discussed below.

In terms of the LIMA model, the mass contained in the equatorial belt is probably $<10^{-10}$ M $_{\odot}$, resulting in a much lower inertia than the WD as a whole. The period-luminosity relationship can be explained by the inner radius of the disc moving further inward as the accretion rate increases, yielding a shorter Keplerian period for the equatorial belt (spin-up). As the accretion rate decreases in the final stages of the outburst, the inner disc boundary moves outward quickly, resulting in a fast outward centrifugal acceleration of the wound-up belt. Gas is now propelled outwards from the belt, removing angular momentum and thus rapidly spinning down the belt. The result is an equatorial belt with a longer Keplerian period and larger associated radius.

Magnetic field lines connecting the belt and the inner disc will be wound up, until reconnection occurs after differential rotation of $\sim 2\pi$, relieving the tension in the field lines (see Warner and Woudt, 2002 and references therein). The observed DNO period “jumps” superimposed on the more gradual variability discussed above are too rapid to also be explained by the spin-up and spin-down of the belt. Latitudinal variations in rotational period may be present in the belt, with the period increasing at higher latitudes towards the rotation period of the underlying primary. MHD turbulence caused by reconnection events may cause so-called “accretion curtains”,

from which gas can be fed through accretion arcs at different latitudes. This could explain the discontinuous jumps in the period.

In a practical sense, many accretion arcs would usually be involved simultaneously, and if they are spread out over different latitudes, a coherent oscillation may not be observable. The detection of DNOs may only be because of the coincidental domination of one or two of the arcs.

The double DNO periods and the often associated DNO-related QPO period that are related as beat periods, can be explained if the shorter DNO is associated with the equatorial belt and the QPO with a structure such as a thickening in the disc (e.g. at the accretion stream impact) revolving around the primary in a prograde direction. The longer DNO is then caused by the reprocessing of the short DNO beam by the QPO structure, and essentially represents a beat between the two.

The class of CVs described above can therefore be considered as an extension of the class of intermediate polars to weaker magnetic fields. The lpDNOs are also thought to be associated with such systems, but are found to have periods equal to approximately one half of the spin period of the WD itself. Warner (2004) suggested that this connects the lpDNOs directly to the WD spin period, with the factor of 2 being due to magnetically channelled two-pole accretion onto the WD. Because $P_{\text{lpDNO}} \sim 4P_{\text{DNO}}$, this implies that the equatorial belt at the inner disc is rotating at an angular velocity 8 times larger than that of the WD itself.

When DNOs and QPOs are simultaneously observed in the optical and X-ray bands, the modulations in these wavebands are of the same time-scale. The properties of DNOs and QPOs are discussed here with regard to CVs, in which they have been extensively studied. However, because of the great similarity between CVs and supersoft X-ray binaries, similar quasi-coherent signals can be expected from the latter class of objects.

In fact, quasi-periodic optical variations with periods of ~ 1.8 h and ~ 5.5 - 7.4 h respectively have been observed in the galactic supersoft sources RX J0019.8+2156 and MR Vel (Deufel et al., 1999; Schmidtke et al., 2000; McGrath et al., 2001). These variations have been explained by luminous blobs of material passing through the disc, or by variable structures in the accretion disc rim. The latter interpretation is favoured by the observed periods being comparable to the Keplerian periods at the outer edge of the disc.

2.8 Outflows in white dwarf binaries

Evidence for the presence of outflows has been found in several supersoft X-ray binaries. The galactic sources RX J0019.8+2156 and MR Vel, as well as the LMC source RX J0513.9-6951, exhibit sharp blue- and redshifted “satellite” features mirrored on the main Balmer emission lines (Southwell et al., 1996a; Becker et al., 1998; Tomov et al., 1998; Motch, 1998). Their velocity displacements from the central emission component are ~ 885 , ~ 3800 and ~ 5200 km s $^{-1}$ respectively. Although the latter two sources are thought to have low inclinations, eclipse-like minima in the lightcurve of RX J0019.8+2156 indicate a much higher inclination, and if the outflows are approximately perpendicular to the accretion disc, a projected velocity of ~ 885 km s $^{-1}$ then indicates an actual jet velocity of several thousand km s $^{-1}$ (Tomov et al., 1998 and references therein). The source CAL 83 in the LMC exhibits broad, variable Doppler shifted wing structures that should also be associated with an outflow, albeit a “wider” one with a larger collimation angle.

The Balmer lines in all four of these SSSs have been observed to have blueshifted absorption features or “P Cygni” profiles (see §2.8.3 below), serving as additional evidence for mass loss. Bearda et al. (2002) also found P Cyg structures in the O VIII (Ly β) and Fe XVII lines in the X-ray spectrum of MR Vel, but with a lower velocity of ~ 1500 km s $^{-1}$.

Evidence for outflows has also been found in CVs, primarily in the form of P Cyg profiles in spectral lines in the ultraviolet waveband (see e.g. Froning, 2005 for a review), similar to those observed from expanding radiation-driven winds in O stars and Wolf-Rayet stars. The wind mass-loss rates can vary from $\lesssim 1\%$ to several times 10% of the mass accretion rate in the disc. The blue edge velocities of the absorption components are typically ~ 5000 km s $^{-1}$, providing a lower limit to the terminal velocity of the wind in question.

The escape velocity from a region close to a WD is

$$v_{\text{esc}} = 5150 \left(\frac{M_1}{M_\odot} \right)^{1/2} \left(\frac{R}{10^9 \text{ cm}} \right) \text{ km s}^{-1}, \quad (2.98)$$

illustrating that the high-velocity outflows in CVs and SSSs should originate near the WD. Actually, this excellent agreement between the WD escape velocity and the observed outflow velocities in SSSs serve as additional evidence that these systems do indeed contain accreting WDs.

Outflows are generally divided into two classes: winds and jets. The distinction is not always clear, but “jets” generally refer to narrow or well-collimated outflows

propagating in a direction more or less perpendicular to the disc surface, with terminal velocities approximately equal to the escape velocity of the central object. Winds, on the other hand, are not so well collimated, and exhibit velocities ranging from approximately $0.1v_{\text{esc}}$ to v_{esc} (Murray, 2002).

Early models of radiation-driven winds in CVs involved radial winds (i.e. directed close to the orbital plane), but this was not consistent with observational evidence (see Froning, 2005 and references therein). One of the problems is that, in CVs with a high inclination where the disc is viewed more edge-on, the blueshifted absorption seen in the lower-inclination systems is not visible. This argues in favour of an outflow more perpendicular to the orbital plane, therefore modern models of radiation-driven CV outflows involve bipolar or “biconical” winds, which are also rotating.

However, even though evidence for *winds* have been found in CVs, evidence for *collimated jets* have been notoriously elusive in these systems (e.g. Knigge and Livio, 1998; Hillwig, Livio and Honeycutt, 2004; Soker and Lasota, 2004). This led Soker and Lasota (2004) to derive a lower limit for the accretion rate above which jets can be driven from accretion discs around WDs: $\dot{m}_{\text{acc}} \gtrsim 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$. They note, however, that the relevant uncertainties may cause this limit to be overestimated, and that it might be as low as $\gtrsim (1-3) \times 10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}$.

We note that bipolar $\text{H}\alpha$ satellite features have been discovered in the recurrent nova T Pyx, but this source is similar to SSSs in the sense that hydrogen burning is still taking place on the WD surface as part of a slow nova outburst, even though its accretion rate is lower than in SSSs (Shahbaz et al., 1997). Bipolar collimated outflows have also been observed recently in RS Oph (Sokoloski et al., 2008) and V445 Pup (Woudt, Steeghs, Karovska, Warner, Groot, Nelemans, Roelofs, Marsh, Nagayama, Smits and O’Brien, 2009).

It is very interesting to consider CAL 83 and RXJ0513 in the context of the limit imposed by Soker and Lasota (2004). As discussed in Chapter 3, CAL 83 is expected to have a lower accretion rate than RXJ0513, and it is very likely that $\dot{m}_{\text{acc}}$ might be slightly below this limit for the former and slightly above for the latter. Indeed, collimated jets are observed in RXJ0513 and not in CAL 83, although both of these systems exhibit evidence of winds.

Broadly speaking, the outflows discussed above can be driven by one of two mechanisms: the outward radiation pressure, or by magnetic fields. These concepts are briefly explained below; for a review, see e.g. Proga (2005) and references therein.

2.8.1 Radiation-driven winds

Atoms close to the primary star will be subject to the gravitational attraction of the primary, but will also be bombarded by high-energy photons pushing them in the opposite direction. The transfer of momentum from the photons to the material can cause the material to be blown away from the primary in what is known as a *radiation-driven wind*.

One type of radiation-driven wind involves momentum transfer by electron scattering by photons in the radiation field. This is often referred to as a *continuum-driven* wind. However, Castor, Abbott and Klein (1975) also showed that the transfer of momentum via line opacity can be several orders of magnitudes stronger than that due to electron scattering by the radiation continuum. Line-driven winds have been shown to be very effective in producing the outflows observed in CVs (Proga, 2005).

One must keep in mind, though, that although SSSs are related to CVs, the WD surface temperature is much higher due to the nuclear burning taking place, yielding a much higher degree of ionization of the surrounding material. A highly ionized plasma will have a very low line opacity, making the line-driven mechanism very inefficient. Also, the luminosity of SSSs is approximately 5 to 7 orders of magnitude higher than that of CVs, meaning that the continuum-driven mechanism will be much more effective than in CVs.

By considering only the electron scattering opacity, Hachiya, Tajima and Fukue (1998) calculated the properties of winds in supersoft X-ray sources, driven by radiation from an *active* accretion disc, as well as from the boundary layer. It was assumed that the disc is optically thick and geometrically thin. They found that these winds can be driven at velocities of $\sim 0.6v_{\text{esc}}$, and that the trajectories of winds depend on the geometry of emission from the central star. When the radiative zone is a thin “belt” concentrated at the equator, the angle between the wind and the normal to the disc is $\sim 45^\circ$. The larger the area of the radiative zone, the more the winds tend to blow in the radial direction, i.e. in the plane of the disc.

Based on the shapes of their lightcurves and their emission line widths, the SSSs CAL 83, RX J0513.9-6951 and MR Vel are expected to have low inclinations (e.g. Cowley et al., 1998), and thus their high projected outflow velocities support a bipolar- and not a radial outflow structure. A solution to this discrepancy was presented by Fukue and Hachiya (1999), who extended the study to continuum-driven winds from *passive* discs. They considered three radiation components: the direct emission from the central star, the emission from the central star reflected once from the disc, and the radiation from the central star that is absorbed and reprocessed

in the disc. Their calculations revealed that: (1) gas originating from the surface of the central star will be blown off in a bipolar wind, and (2) gas originating from the inner disc will form an equatorial outflow, “skimming” outward over the surface of the accretion disc. Therefore a radiation-driven bipolar wind is indeed a probable occurrence in SSSs.

2.8.2 Magnetospheric outflows

In §2.5.2, it was explained how a WD with a strong magnetic field can channel the accreting flow towards its polecaps. The WD and its magnetosphere is rotating with angular velocity Ω_* . Since the accreting material must start flowing along the magnetic field lines at R_M , the value of Ω_* must be smaller than $\Omega_K(R_M)$ for steady accretion to occur, i.e. the WD must be a slow rotator, with $\omega_* < 1$ (refer to §2.6.1). On the other hand, if $\omega_* \sim 1$, the particles trying to accrete onto the fast-rotating magnetosphere will be repelled by this “centrifugal barrier”, and be propelled outwards. Put in another way: if the corotation radius lies outside the Alfvén radius ($R_\Omega > R_M$), steady accretion can take place, otherwise accreting material will be flung outwards.

In particular, Blandford and Payne (1982) illustrated that an effective wind can be driven from an accretion disc by the centrifugal force if the angle between the poloidal component of the magnetic field and the plane of the disc is less than 60° . This can occur, for example, in a supersoft X-ray binary accreting at $\sim 10^{-7} \text{ M}_\odot \text{ yr}^{-1}$, with the high influx of material at the disc-magnetosphere boundary “compressing” the magnetic field and bending the magnetosphere inwards (see Fig. 2.17). These are often referred to as *magnetocentrifugal winds*, and generally require that the poloidal field strength should be at least comparable to the toroidal field strength. At large distances from the disc, the toroidal component of the magnetic field becomes important and acts to collimate the wind into a set of bipolar jets perpendicular to the disc.

These types of outflows have been the object of quite extensive numerical studies, which have painted a very promising picture regarding the ejection of magnetically driven winds under a variety of circumstances. Some interesting recent results include the following: The half-opening angles of the jets collimated by magnetic fields are typically in the $4\text{--}20^\circ$ range, depending on the degree of flow magnetisation (Lii, Romanova and Lovelace, 2011). Importantly, it has appeared that the formation of magneto-centrifugally accelerated winds from an accretion disc threaded by a weak vertical magnetic field is in fact a natural consequence of the existence of the magnetorotational instability, the latter being an almost essential element for disc accretion (Fromang et al., 2013). Large-scale, unstable magnetorotational instabilities in an accretion disc in which the magnetisation is moderately subthermal will

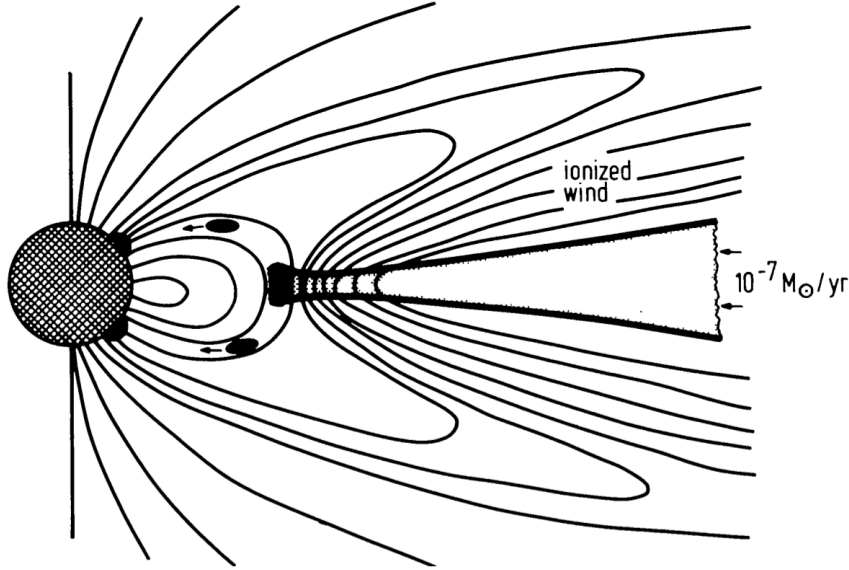


Fig. 2.17: The geometry of a magnetocentrifugal wind from a magnetized accreting star. (This diagram was adopted from Camenzind, 1990, Fig. 12, who used it to illustrate the magnetic field geometry of an accreting T Tauri star, but the same field geometry can be applied to the high-rate mass accretion onto WDs in supersoft X-ray binaries.)

spontaneously produce magnetically driven outflows in the non-linear regime (Lesur, Ferreira and Ogilvie, 2013). Angular momentum is transferred to the bending field lines (found in the Blandford and Payne, 1982 model), and is then released to the accelerated material in the wind.

If an accretion disc is coupled with an initially vertical magnetic field, the magnetorotational instability will lead to a growth in the strength of the toroidal magnetic field as the field lines are wound up by the rotating disc (Proga, 2005 and references therein). Under these circumstances, the magnetic pressure of the substantial toroidal field can drive a wind as well.

In summary: the high-velocity winds observed in SSSs originate from near the WD surface. They may be continuum radiation-driven winds, but the highly probable presence of magnetic fields in the inner disc region indicates that magnetic field activity may also be a significant outflow-driving mechanism.

2.8.3 The P Cygni profile as an outflow signature

In some astronomical sources, a blueshifted absorption feature is present in addition to the main emission line component. This has become known as a P Cyg profile, named after the blue variable star P Cygni (e.g. Hellier, 2001, pp. 101-103). In P Cyg, as well as other similar stars, these profiles are produced by a stellar wind from the star, as shown in Fig. 2.18. A part of the wind is hidden from the observer by the central star, and therefore not visible. The regions of the optically thin wind

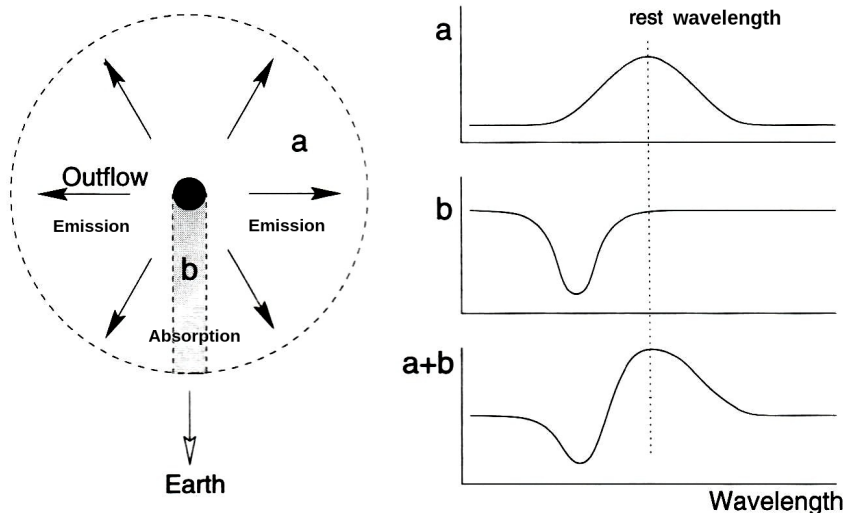


Fig. 2.18: The formation of a P Cyg profile in a wind blowing from the central star. (Adapted from Hellier, 2001, Fig. 7.5, p. 102.)

seen on each side of the star contain material with a range of radial velocities, and yield a broadened emission line in the spectrum.

The part of the wind producing the largest blueshift is the material moving directly towards the observer. However, this region will produce an absorption component, as it will absorb the continuum emission from the side of the star facing the observer, and re-emit it in all directions. The combined result is an emission line with a blueshifted absorption component (see Fig. 2.18). Many other sources that undergo mass loss with a more complex geometry also exhibit P Cyg profiles, for example winds from the boundary layer or inner disc in WD binaries, therefore this structure has become a very useful signature and diagnostic when studying outflows (e.g. Drew, 1995).

2.9 Classification and evolution of supersoft X-ray sources

In addition to the close binary supersoft sources described by Van den Heuvel et al. (1992), there are numerous other types of sources that can also produce supersoft X-rays. Di Stefano and Nelson (1996) put forward a classification scheme describing four different classes of hydrogen-burning accreting WD SSSs: CVs, CBSSs, WBSSs (wide-binary supersoft sources) and SySSs (symbiotic supersoft sources). This is shown in Table 2.1.

In §2.6.3, it was explained how nova outbursts in CVs can lead to supersoft X-ray emission. As suggested by Di Stefano (1996), the so-called WBSSs may also exhibit SSS behaviour. These are systems with a donor that was initially more evolved than the donor in CBSSs, and where an accretion rate in the steady burning strip

Table 2.1: The SSS classification scheme of Di Stefano and Nelson (1996) for hydrogen-burning C-O white dwarfs. (Adopted from Odendaal, 2012, as adapted from Di Stefano and Nelson, 1996, Table 1.)

Type	Mass transfer mechanism	$\dot{m}$ ($M_{\odot} \text{ yr}^{-1}$)	P_{orb} (days)	Steady (S) / Recurrent (R)	Mass ejection	Well-known examples
CV	Magnetic braking/ gravitational radiation	$< \sim 10^{-8}$	$< \sim 0.2$	R	Nova	Nova LMC 1995 (CN) U Sco (RN) SMC 13?
CBSS	Thermal time-scale readjustment of donor	$> \sim 10^{-7}$	$\sim 0.2 - 3.0$	S	Winds (for $\dot{m} > \dot{m}_{\text{cr}}$)	CAL 83 CAL 87 RX J0513.9-6951 MR Vel
WBSS	Nuclear evolution of donor	$< \sim 10^{-7}$ $> \sim 10^{-7}$	$\sim 0.2 \times 10^2$ $\sim 3.0 \times 10^2$	R S	Nova Winds (for $\dot{m} > \dot{m}_{\text{cr}}$)	
SySS	Stellar winds or Roche lobe overflow from evolved donor	$< \sim 10^{-7}$ $> \sim 10^{-7}$	$\sim 3.0 \times 10^2$ $\sim 10^2$	R S	Nova Winds	RR Tel AG Dra RX J0048.4-7332
		$< \sim 10^{-7}$	$\sim 10^2$	R	Nova	

is sustained by Roche lobe overflow due to the nuclear evolution of the donor. Symbiotic stars are interacting binary stars with very long periods, usually consisting of a hot WD and a red giant companion, and several of these systems have also been observed as SSSs.

The secondary star in CBSSs is assumed to be slightly evolved, i.e. between zero-age main-sequence and terminal-age main-sequence or even in the subgiant phase. Although estimations of secondary masses have been obtained with indirect methods (Kahabka and van den Heuvel, 2006 and references therein), the donors are generally too faint in comparison with the bright accretion disc, and their spectral types can not be determined.

Some SSSs consisting of single stars have also been observed, and are believed to consist of evolved stars that are in the pre-WD phase, like PG 1159 stars and the nuclei of planetary nebulae (e.g. Kahabka and van den Heuvel, 2006 and references therein). Yungelson et al. (1996) also discussed the possibility of systems with He donors, and also double degenerate systems, as SSS candidates.

The origin of CBSSs is closely related to that of CVs, with the main difference being that the eventual donor is more massive than the WD that is formed. These close binary supersoft sources initially form through common envelope (CE) evolution of wide binary systems containing e.g. a $\sim 7 M_{\odot}$ red giant with a degenerate C-O or O-Ne-Mg core, and a main-sequence star with mass in the $1.5\text{--}3.0 M_{\odot}$ range (e.g. Kahabka and van den Heuvel, 1997; Bitzaraki et al., 2004). The symbiotic systems also

form through CE evolution, but the initial binary orbit is so wide that the originally more massive red giant evolves into a WD without noticeable interaction with the companion. Mass transfer only starts when the companion evolves into a red giant.

A WD is supported against its own gravitation by electron degeneracy pressure. The maximum possible mass for an object supported by electron degeneracy is $\approx 1.4 M_{\odot}$, which is known as the *Chandrasekhar limit*. When a massive WD approaches the Chandrasekhar limit by accretion, the ignition of nuclear fuel in its core can cause the ignition of a runaway explosive event called a Type Ia supernova (SN Ia), completely destroying the WD. WDs in steady SSSs accreting without substantial loss of accreted material are considered to be promising candidates for SN Ia progenitors (e.g. Nomoto et al., 2007; Podsiadlowski, 2010).

Alternatively, before the onset of a SN Ia, the WD can collapse to a neutron star through inverse beta decay, a process known as accretion-induced collapse (AIC) (e.g. Metzger et al., 2009). SN Ia and AIC are therefore two competing processes, and the fate of the white dwarf will depend on which of the two processes commences first (see also Di Stefano et al., 2010). An accreting C-O WD may tend to end in a SN Ia, and an O-Ne-Mg WD in an AIC.

Another possibility is that supersoft sources may evolve to become cataclysmic variables. Meintjes (2002) and Schenker et al. (2002) explained the long orbital period and short white dwarf spin period (~ 33 s) of the CV AE Aqr by the possibility that the source has passed through a supersoft phase in its recent past. It is thought that a significant fraction of the current cataclysmic variables may have formed in this way.

The discussion presented here serves to summarize most of the relevant properties of these systems, as well as to provide a guide for relevant observational signatures that can be utilized to decode the secrets of these fascinating sources. The following chapter will focus on the individual properties of the supersoft X-ray binaries forming part of this study.

Chapter 3

Literature review of target sample

The three supersoft X-ray binaries focussed on in this study are CAL 83 and RX J0513.9-6951 (RXJ0513) in the LMC, and RX J0925.7-4758 (MR Vel) in the Milky Way. These three southern sources were selected because they have several properties in common that make them very good candidates for the current study: Each source contains a massive WD accreting through an accretion disc, which may have been spun up to a short rotation period by disc torques during an extensive period of high-rate mass accretion. All three exhibit evidence for outflows in the form of Doppler shifted emission features and P Cyg absorption features. The inclinations of all three systems are quite low, providing a relatively unobstructed view of the approaching outflow component, as well as the central white dwarf and the surrounding accretion disc.

In this chapter, the currently known properties of these sources are presented in §3.1, §3.2 and §3.3 respectively. Each of these sections is subdivided into four subsections containing the target characteristics in different wavebands: X-ray, ultraviolet, optical and longer wavelengths (infrared to radio). The exception is MR Vel, for which no ultraviolet observations have been reported to date.

The observational investigation presented in this thesis is focussed on the X-ray and optical variability related to accretion. The goal of this chapter is therefore to provide a detailed review of the published observational work and relevant interpretations for each target. Within each subsection, the literature is presented more or less chronologically in the form of an historical overview, illustrating how the increasing observational evidence over the years has led to our current understanding of these close binary supersoft sources. This will provide the reader with a comprehensive background regarding the known source properties and their variability, before proceeding to the new observational results presented in the subsequent chapters.

3.1 CAL 83

3.1.1 X-ray

CAL 83 was source number 83 in the Columbia Astrophysics Laboratory survey of the Large Magellanic Cloud (LMC), using the Imaging Proportional Counter (IPC) aboard the *Einstein Observatory* (Long et al., 1981; Brown et al., 1994). It was the brightest X-ray point source discovered in this survey that is not associated with a Galactic foreground star or a background AGN. The parameters resulting from a blackbody fit to the IPC data by Brown et al. (1994) are given in Table 3.1, together with other X-ray model parameters for CAL 83 from the literature. The LMC distance of $d \sim 50$ kpc (Pietrzyński et al., 2013) is usually adopted for the LMC SSSs.

From *EXOSAT* observations in November 1983, Pakull, Ilovaisky and Chevalier (1985) also confirmed the supersoft nature of CAL 83. They remarked on the low X-ray to optical flux ratio and, considering the source as an LMXB, ascribed it to a possible geometry where the X-ray emitting compact object is obscured from direct view by an accretion disc, with a small fraction of the X-ray flux scattered into the line of sight by the accretion disc corona. CAL 83 was observed again with the *EXOSAT Observatory* on 18 December 1985, and the photon spectral index was estimated to be ~ 3 in the 0.05-2.0 keV band (Crampton et al., 1987). The latter paper also remarked on the soft spectrum being reminiscent of some black hole candidates.

During the calibration phase of the Röntgensatellit (*ROSAT*) in June 1990, and also during the *ROSAT* all-sky survey in 1990-1991, CAL 83 was one of the sources observed (Greiner et al., 1991; Schmidtke et al., 1994; Trümper et al., 1991). The blackbody model fit parameters given by Greiner et al. (1991) are presented in Table 3.1.

Gänsicke et al. (1998), on the other hand, fitted LTE WD model atmosphere spectra to the *ROSAT* PSPC spectrum, and obtained a lower value for the bolometric luminosity than obtained with the blackbody model. (Refer to §2.6.5 for an introduction to WD atmosphere models). They obtained the best fit values for kT_{eff} , N_{H} and $(R/d)^2$ for $\log g = 8$, as well as $\log g = 9$. In both cases, the best fit column density is $7.0 \times 10^{20} \text{ cm}^{-2}$. Kahabka (1998) also performed LTE and NLTE WD model atmosphere fits to the *ROSAT* data for $\log g = 9$ (see Fig. 3.1). The best fit parameters from these papers are listed in Table 3.1.

Greiner et al. (1991) remarked on the fact that the flat shape of the optical-to-UV spectrum indicates that the source is an interacting binary system. However, they argued that the extreme softness of the X-ray spectrum can not be explained well

Table 3.1: A summary of the X-ray properties of CAL 83, as obtained from published models fitted to the X-ray spectra. BB stands for blackbody and LTE refers to WD model atmosphere in local thermodynamic equilibrium, while non-LTE (NLTE) models allow departures from LTE. For more details regarding the particular WD atmosphere models, please refer to the relevant articles. The distance d to the source was assumed to be 48 kpc by Brown et al. (1994), while all the other authors used a value of 50 kpc (this was not explicitly stated by Kahabka, 1998, but we assume a similar value was used). Where available, the χ^2 -value and degrees of freedom (dof) of the fit are also given.

Facility	Model	$\log g$ (cgs)	kT_{eff} (eV)	N_{H} (10^{20} cm^{-2})	R (10^8 cm)	L_{bol} (10^{37} erg s^{-1})	χ^2 /dof	M_{wd} ($M_{\odot}$)
<i>Einstein</i> ^[1]	BB	–	~ 11	7.1 ^a	590	65	7.7/7	
<i>ROSAT</i> ^[2]	BB	–	42-48	7 ^a	~ 12	~ 7		
<i>ROSAT</i> ^[3]	LTE	8 ^a	42.6 ± 1.0	7	$5.86^{+0.72}_{-1.06}$	$1.46^{+0.33}_{-0.51}$		
<i>ROSAT</i> ^[3]	LTE	9 ^a	49.1 ± 1.5	7	$4.47^{+0.30}_{-1.12}$	$1.50^{+0.08}_{-0.73}$		
<i>ROSAT</i> ^[4]	LTE	9 ^a	$44.5^{+1.1}_{-1.0}$	$8.87^{+1.55}_{-1.30}$	10.0-11.4	$5.81^{+0.13}_{-0.20}$	42.5/36	
<i>ROSAT</i> ^[4]	NLTE	9 ^a	$\lesssim 34.7$	$6.07^{+\geq 0.33}_{-\geq 0.47}$	$\gtrsim 10.3$	$\gtrsim 2.0$	53.5/36	
<i>ROSAT</i> ^[5]	LTE ^b	8.0-8.25	43.4 ± 1.5	6.33 ^a	7-9	2.8-3.1	17.3/9	0.6-0.8 ^c 1.2-1.3 ^d
<i>BeppoSAX</i> ^[6]	BB	–	46.4 ± 1.4	6.5 ^a	8.26 ± 0.30	2.49 ± 0.19	17.5/23	
<i>BeppoSAX</i> ^[6]	LTE	9.0 ^a	$44.0^{+6.3}_{-0.6}$	$9.5^{+0.3}_{-5.2}$	$10.5^{+0.8}_{-7.6}$	$5.4^{+0.6}_{-4.7}$	22.4/22	
<i>BeppoSAX</i> ^[6]	NLTE	8.0 ^a	$31.7^{+0.5}_{-3.8}$	$5.4^{+4.0}_{-2.2}$	11^{+32}_{-7}	$1.6^{+7.9}_{-1.0}$	20.2/22	
<i>BeppoSAX</i> ^[6]	NLTE	8.45 ^a	32.6 ± 0.7	6.5 ^a	$12.5^{+0.8}_{-0.5}$	$2.30^{+0.10}_{-0.05}$	17.5/23	
<i>BeppoSAX</i> ^[6]	NLTE	8.5 ^a	32.4 ± 1.2	$6.8^{+1.4}_{-1.0}$	13.5^{+6}_{-3}	$2.6^{+2.0}_{-0.8}$	17.4/22	
<i>BeppoSAX</i> ^[6]	NLTE	9.0 ^a	$34.15^{+2.8}_{-0.35}$	$7.8^{+0.5}_{-2.3}$	14^{+27}_{-7}	$3.4^{+0.7}_{-2.1}$	17.8/22	
<i>Chandra</i> , <i>XMM</i> ^[7]	NLTE ^b	8.5 ± 0.1	47.4 ± 2.2	6.5 ^a	7.0 ± 0.7	3.4 ± 1.1		1.3 ± 0.3
<i>XMM</i> ^{[8],e}	BB	–	18-37	6.5-24.3	16.0-164.5	2-118	23/36-235/82	

[1] Brown et al. (1994)

[2] Greiner et al. (1991), see also the comments of Van den Heuvel et al. (1992)

[3] Gänsicke et al. (1998)

[4] Kahabka (1998)

[5] Suleimanov and Ibragimov (2003)

[6] Parmar et al. (1998)

[7] Lanz et al. (2005)

[8] Rajoelimanana et al. (2013)

^a Fixed during fit.

^b Line-blanketed.

^c Assuming photospheric radius is equal to radius of a cool degenerate WD.

^d Assuming photospheric radius is equal to twice the radius of a cool degenerate WD.

^e Including error ranges.

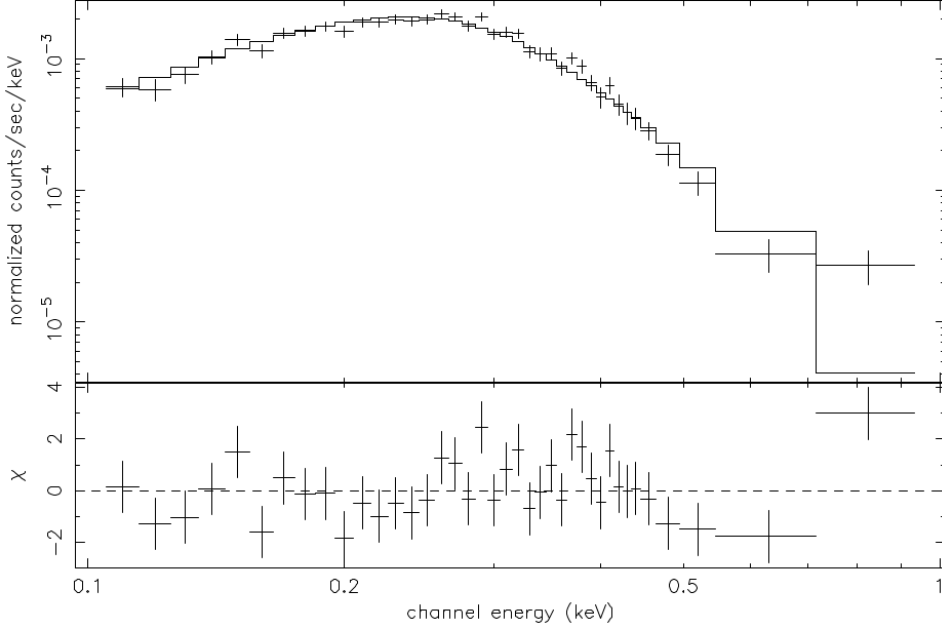


Fig. 3.1: NLTE WD model atmosphere fit to *ROSAT* PSPC spectrum. (Adopted from Kahabka, 1998, Fig. 3.)

by the model of an LMXB with an accretion disc corona. They offer an alternative LMXB interpretation, which involves a compact object undergoing super-Eddington accretion, with the piled-up material forming an ionized cocoon around the compact object that is optically thick for Compton scattering. The result is a softer spectrum with a photospheric radius comparable to that of a WD.

Greiner et al. (1991) did suggest the possibility of CAL 83 being a nova, but, based on the observational evidence, the outburst must then have been either before 1973 or between 1975 and the *Einstein* survey. However, then one would expect a decline in X-ray flux between the *Einstein* and *ROSAT* observations, which was not observed.

Soon after that, Van den Heuvel et al. (1992) proposed steady nuclear burning on the surface of a WD accreting at a rate close to the Eddington limit, rather than a nova outburst, as the probable cause of the supersoft X-ray emission from CAL 83. This has since been the favoured model for CAL 83 and many other SSSs (see §2.1). According to their model, the WD in CAL 83 has a mass of $\gtrsim 0.8 M_{\odot}$ and is accreting at a rate of $\sim 4 \times 10^{-7} M_{\odot} \text{ yr}^{-1}$ from an (evolved) donor star with a mass in the $1.5\text{--}2 M_{\odot}$ range. The system is expected to have an X-ray lifetime of $\sim 10^7$ years, which is typically of the order of the thermal time-scale of the secondary star.

On 28 March and 21 April 1996, CAL 83 was observed again in X-rays by *ROSAT*. However, surprisingly, another observation on 28 April 1996 found CAL 83 to be in

an X-ray off-state (Kahabka, 1996b, 1997). This non-detection represented a reduction in X-ray flux by a factor of >43 over 21 days. The associated decay time is $\lesssim 6$ d for exponential decay, and $\lesssim 20$ d for linear decay. An observation on 6 August 1996 showed the source at the same flux level as before the off-state (Kahabka et al., 1996).

The X-ray off-state can either be caused by a decrease in the bolometric luminosity, or by a decrease in the source temperature (shifting the emission towards longer wavelengths), or both. Alcock et al. (1997) suggested two possible models as an explanation of the X-ray off-state: (1) A cessation in nuclear burning could cause the WD envelope to cool and the luminosity to decrease, which would occur if the accretion rate decreases to a value below the value of $\dot{m}_{\text{std}}$ required for steady nuclear burning (as explained in §2.6.3). (2) An expansion in the WD photosphere, like at the beginning of a nova outburst, would increase the photospheric radius and therefore decrease the effective temperature, shifting the peak of emission into the UV/optical wavebands.

Alcock et al. (1997) compared the changes in X-ray flux for the *ROSAT* observations to the MAssive Compact Halo Objects (MACHO) Project lightcurve of CAL 83, and although the data points are quite sparse, the general conclusion was that the source undergoes an optical minimum ~ 10 -15 d before an X-ray minimum (Fig. 3.2). This would be in agreement with model (1) above, because as the accretion rate decreases, the optical luminosity from the disc will decrease, where-after the steady nuclear burning will cease, and the X-rays will turn off. When the accretion rate increases again to the original value, the disc and X-ray luminosity will return to the pre-minimum state. As the steady burning strip for WDs is relatively narrow (Fig. 2.12), a decrease in the accretion rate is very likely to be able to turn off the steady burning. Such a decrease in the accretion rate may be caused by star spots on the surface of the secondary star that become aligned with the L_1 point, as in the VY Scl stars (Livio and Pringle, 1994).

Upon the cessation of nuclear burning, the X-ray luminosity will decline. Prialnik and Kovetz (1995) considered the parameter $t_{3\text{bol}}$, the time it takes for the bolometric luminosity of a WD with surface nuclear burning to decline by 3 magnitudes, and calculated an extended grid of values. Alcock et al. (1997) considered their results, and found that, under model (1), for an X-ray decay time of $\lesssim 20$ d, the WD in CAL 83 must be massive, i.e. $M_{\text{wd}} \gtrsim 1.3 M_{\odot}$. (A decay time of 4.3 d would imply $M_{\text{wd}} = 1.4 M_{\odot}$.)

Model (2), on the other hand, would imply an anti-correlation between X-ray and optical flux, i.e. a decline in the X-rays would be accompanied by a rise in the optical, which was not the favoured interpretation of Alcock et al. (1997) at the time. How-

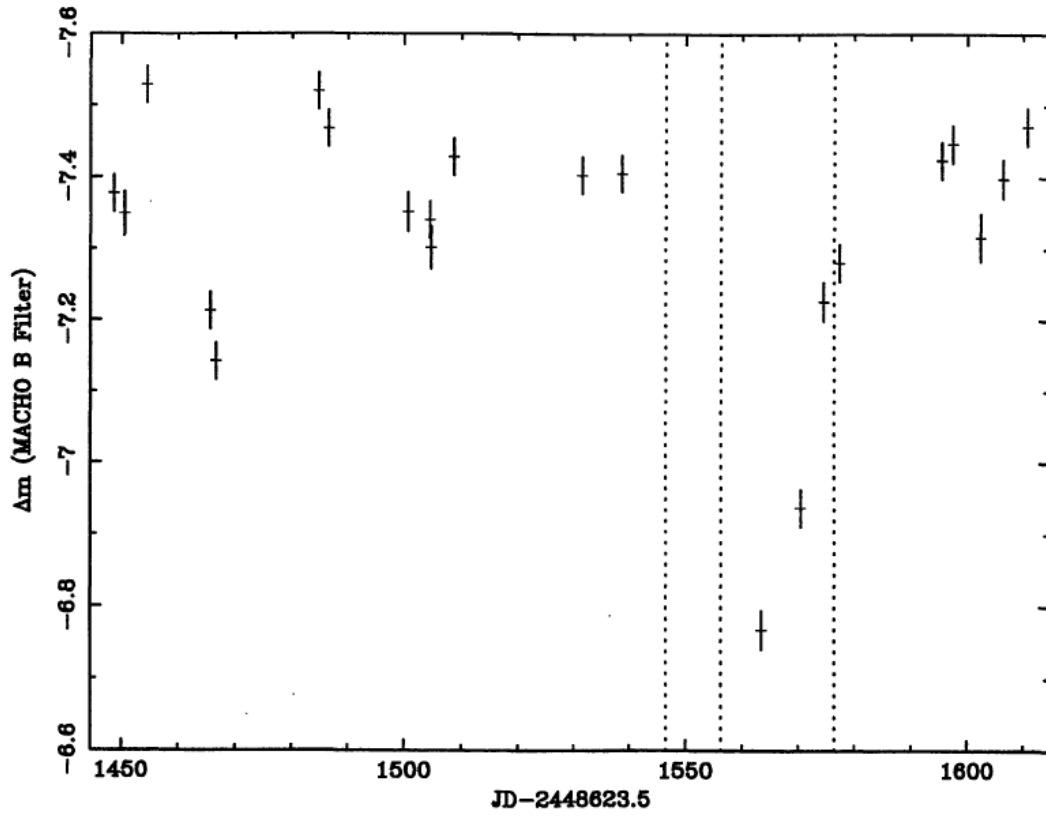


Fig. 3.2: MACHO Project optical lightcurve of CAL 83, with the times of the *ROSAT* observations on 28 March, 7 April and 28 April 1996 overplotted as dotted lines. The first two *ROSAT* observations were during an X-ray on-state, and the third during an X-ray off-state. (Adopted from Alcock et al., 1997, Fig. 1.)

ever, Kahabka (1998) favoured this model, and by estimating the expected *ROSAT* count rate changes that would be accompanied by source temperature changes, they found that the X-ray off-state could be explained by a drop in temperature from 39 eV to 28 eV. From considerations of the turn-off and turn-on time-scales as determined from all the *ROSAT* data (obtained with the Position Sensitive Proportional Counters, or PSPC, and the High Resolution Imager, or HRI), he constrained the WD mass to $\gtrsim 1.2 M_{\odot}$.

Kahabka (1998) proposed an increase in accretion rate as the cause for the temperature decrease. He proposed that the X-ray off-state may have been triggered by a “mass transfer wave” travelling from the donor towards the WD, i.e. a mass transfer enhancement followed by a mass transfer deficit, possibly related to magnetic star spot activity on the donor star. After a time that is mainly determined by the viscous time-scale of the disc, the wave arrives at the WD envelope and triggers an envelope expansion, leading to a drop in X-ray flux as the emission shifts to longer wavelengths. Hereafter, the density wake will cause the envelope to settle back again, and X-ray emission is restored.

From a re-analysis of the *ROSAT* PSPC data with line-blanketed LTE WD atmosphere models, and a comparison with theoretical relationships between the properties of WDs with hydrogen shell burning, Suleimanov and Ibragimov (2003) obtained new estimates for the source parameters of CAL 83. These are also listed in Table 3.1.

CAL 83 was observed with the Low-Energy Concentrator Spectrometer (LECS) on the *BeppoSAX* satellite on 7-9 March 1997 (Parmar et al., 1998). These authors fitted blackbody, LTE and NLTE models to the spectrum (see Table 3.1). The source was also detected by the Advanced Satellite for Cosmology and Astrophysics (*ASCA*) on 21/22 June 1997 at a similar intensity as the *ROSAT* observations, and with an effective temperature of 29 ± 8 eV (Dotani et al., 2000).

The first high-resolution (~ 0.05 Å) X-ray spectrum of CAL 83 was obtained with the Reflection Grating Spectrometers (RGS) on board *XMM-Newton* on 23 April 2000 (Paerels et al., 2001, observation ID 0123510101). The spectrum exhibited various absorption features, and it was evident that any smooth continuum spectrum does not represent an accurate model for the spectrum. Paerels et al. illustrated the excellent qualitative resemblance between the CAL 83 spectrum and the emission associated with a hot, line-blanketed WD atmosphere model. This confirmed the accreting WD model for CAL 83.

Greiner and Di Stefano (2002) reported the observation by *Chandra* of a second X-ray off-state on 29-30 November 1999 (obs IDs 77 and 84), with an upper limit of

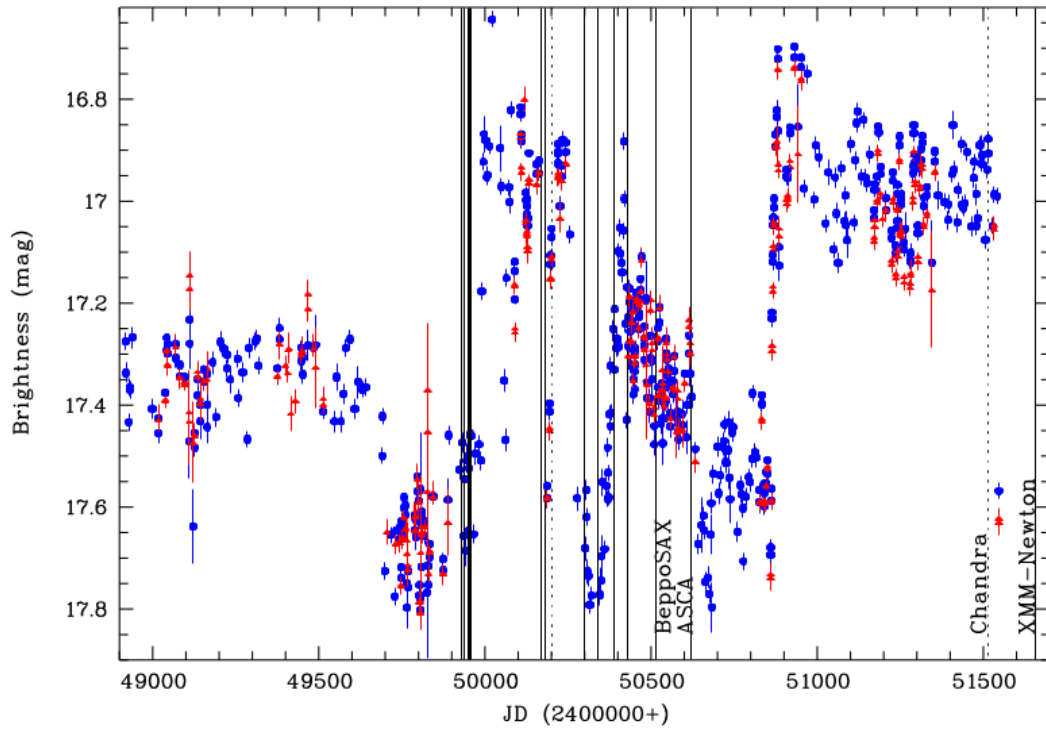


Fig. 3.3: Comparison of the MACHO lightcurve of CAL 83 with its X-ray states. The V-band observations are indicated by the blue filled circles, while the R-band observations are indicated by the red triangles. The solid vertical lines represent times of X-ray on-state observations, while the dashed vertical lines indicate X-ray off-state observations. All the unmarked X-ray observations were performed by *ROSAT*. (Adopted from Greiner and Di Stefano, 2002, Fig. 4.)

$8 \times 10^{34} \text{ erg s}^{-1}$ assuming $kT_{\text{eff}} \sim 40 \text{ eV}$ and $N_{\text{H}} = 6.5 \times 10^{20} \text{ cm}^{-2}$. Greiner and Di Stefano performed a comparison between all the X-ray observations of CAL 83, and its optical lightcurve from MACHO data spanning 1992-1999. The comparison is shown in Fig. 3.3. The magnitude is evidently highly variable, and seems to switch between low, intermediate and high states. When transitioning to a higher state, it seems that the magnitude “overshoots” somewhat.

Greiner and Di Stefano (2002) found that the detected X-ray on-states have all been during the optical low and intermediate states, while the detected X-ray off-states were both during an optical high state. About 30-50 days after each off-state, optical fading was observed. This was not consistent with model (1) of Alcock et al. (1997) that was discussed above, and Greiner and Di Stefano favoured a model related to (2).

An increase in the photospheric radius of the WD (possibly caused by changes in the accretion rate) will shift the supersoft emission towards optical and UV wavelengths, leading to a non-detection in X-rays. However, this would also increase the amount of irradiation of the accretion disc, increasing the amount of reprocessed radiation

from the disc and leading to an even higher optical flux. This means that the variations in the optical can be quite complex. Once the WD photosphere contracts again, the source will be detectable again in X-rays.

Lanz et al. (2005) reported on two new observations of CAL 83 with *Chandra*. One was performed with the High Resolution Camera (HRC-S) and the Low Energy Transmission Grating (LETG) during an X-ray on-state on 15-16 August 2001 (obs ID 1900). The other was during an X-ray off-state on 3-4 October 2001 (obs ID 3402, with an upper limit of $\sim 9.2 \times 10^{34}$ erg s $^{-1}$ for the X-ray luminosity), and constituted the third detection of an X-ray off-state in CAL 83. In the on-state observation, variations of up to 50% of the X-ray flux were observed on short time-scales, approximately <10% of the orbital period of ~ 1 d.

Lanz et al. (2005) also performed a detailed analysis of the on-state LETG spectrum, together with the *XMM-Newton* RGS spectrum initially published by Paerels et al. (2001), by fitting line-blanketed NLTE WD model atmospheres. Their best fit parameters are given in Table 3.1, and the spectra are shown in Fig. 3.4. These authors found the first conclusive spectroscopic evidence that the WD is massive, with a mass in the 1.0-1.4 $M_{\odot}$ range. According to Eq. (2.25), the radius of a cold WD with a mass of 1.3 $M_{\odot}$ is $R_{\text{wd}} = 4.0 \times 10^8$ cm. The photospheric radius of 7×10^8 cm is a factor of ~ 1.8 times larger than R_{wd} , indicating that the WD in CAL 83 has a swollen atmosphere.

The line-blanketed NLTE model fit of Lanz et al. (2005) is the most sophisticated one published for CAL 83 to date, and also the first one applied to *high-resolution* X-ray spectra. Therefore, these authors were able to constrain the true properties of the source to a higher degree of accuracy than was possible for the other fits listed in Table 3.1. We also note that the line-blanketed NLTE fit of Lanz et al., with $R \sim 1.8R_{\text{wd}}$, yields very similar source parameters as the line-blanketed LTE fit of Suleimanov and Ibragimov (2003) with $R \sim 2R_{\text{wd}}$, the latter being associated with $M_{\text{wd}} \sim (1.2-1.3) M_{\odot}$.

Schmidtke and Cowley (2006) undertook detailed timing analyses of *XMM-Newton* observation 0123510101 on 23 April 2000, as well as *Chandra* observation 1900 on 15-16 August 2001. They found a periodic modulation in the *Chandra* lightcurve with a period of 38.4 min, but no periodicities in the *XMM-Newton* data. The 38.4 min modulation has a peak-to-peak amplitude of $\sim 23\%$ of the mean X-ray flux. These authors ascribed the periodicity to non-radial g^+ mode pulsations of the WD, similar to what has been seen in several novae (refer back to the discussion in §2.6.2).

Observations by the *Swift* X-ray Telescope (XRT) in November 2007 showed CAL 83

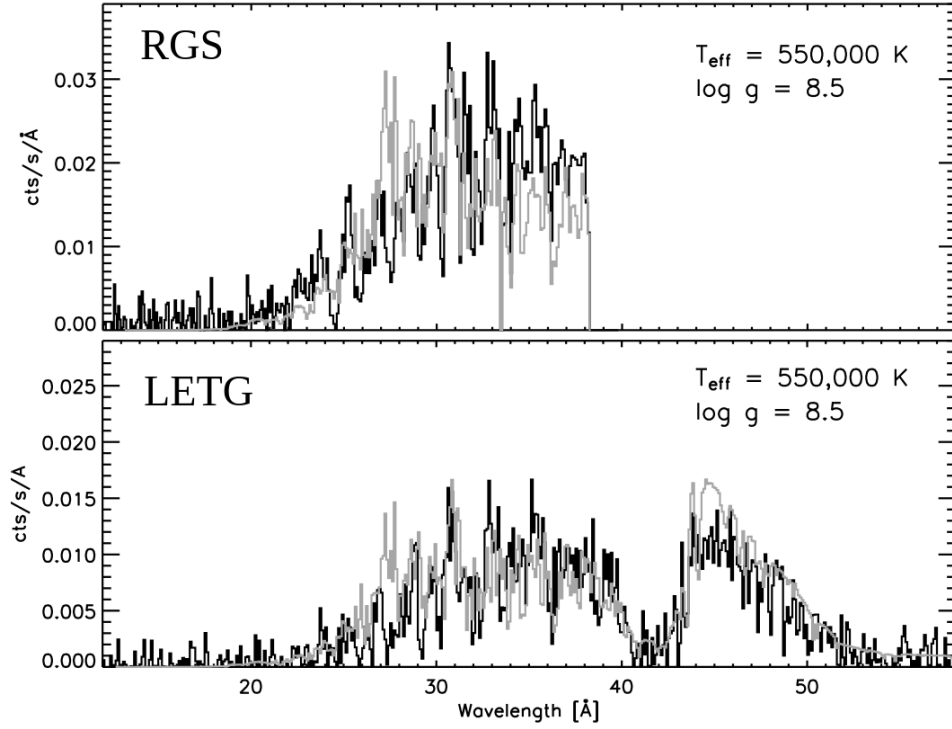


Fig. 3.4: The best fit NLTE model for RGS and LETG spectra. The black lines represent the data, and the grey lines the model. (Adopted from Lanz et al., 2005, Fig. 5 and 6.)

to be in an X-ray on-state. However, after the optical magnitude brightened by 0.5 mag at the end of December 2007, putting the source in its optical high state, *Swift* observed the source in an X-ray off-state on 2 January 2008 (Greiner et al., 2008). The source was still in an off-state by 18 March 2008, but was detected again by *Swift* on 20 March 2008 at 15% of the normal X-ray flux (Lanz, Audard and Walter, 2008). Small & Moderate Aperture Research Telescope System (SMARTS) photometry showed a drop in optical brightness by ~ 0.5 mag about a week before the 20 March 2008 detection. These observations therefore indicate that the X-ray variations lag behind the optical by less than a week. They also confirm the anti-correlation between the X-ray and optical luminosities.

Rajoelimanana et al. (2013) analysed 22 observations of CAL 83 by *XMM-Newton* between May 2007 and May 2009. The epochs of these observations are also indicated on the OGLE-III (Optical Gravitational Lensing Experiment III) lightcurve in Fig. 3.17 in §3.1.3. Four of the observations were during an X-ray on-state, and the rest exhibit variable X-ray flux. These authors distinguished between two distinct X-ray states: a high state where the EPIC pn count rate is ≥ 4.5 counts s^{-1} , and a low state where the X-ray counts are very low or completely undetectable. The X-ray flux is anti-correlated with the optical magnitudes in the OGLE-III lightcurve.

Fixing the column density to $6.5 \times 10^{20} \text{ cm}^{-2}$, Rajoelimanana et al. (2013) fitted

absorbed blackbody models to the EPIC pn spectra of all the observations. They discovered a slight dependence of the effective temperature on the optical state, i.e. ~ 30 eV during optical high states and ~ 35 eV during optical low states. All the findings of Rajoelimanana et al. (2013) are consistent with the photospheric expansion/contraction model.

3.1.2 Ultraviolet

Crampton et al. (1987) reported on International Ultraviolet Explorer (*IUE*) spectra of CAL 83 obtained in November 1984, February 1985 and May 1986. The N v $\lambda 1240$ and He II $\lambda 1640$ emission lines are present in the spectrum, as well as an unidentified feature at $1670 \text{ \AA}$ and possible absorption features near 1260 , 1300 and $1335 \text{ \AA}$. The C IV $\lambda 1550$ line was not detected, even though it could be expected to be present on account of the ionized carbon emission at $\sim 4600 \text{ \AA}$ (see §3.1.3). The continuum temperature is $\sim 20\,000$ K. They found no extended (nebular) flux in the UV continuum or emission lines.

Bianchi and Pakull (1988) also presented additional *IUE* spectra obtained in January and December 1987. They reported an occasion (January 1987) where the flux was higher than in the other observations by a factor of 2 to 3, and the temperature was also higher. Simultaneous European Southern Observatory (ESO) observations in January 1987 confirmed that the optical flux was then also higher (Bianchi and Pakull, 1988). The spectral features were evidently also clearer in the January 1987 spectrum, and included interstellar-like, or possibly circumstellar, absorption features of Si III $\lambda 1260$, O I $\lambda 1300$ and C II $\lambda 1335$, as well as emission features of N v $\lambda 1240$, Si IV $\lambda 1403$ and He II $\lambda 1640$.

From preliminary modelling, Bianchi and Pakull (1988) interpreted the UV and optical flux as originating from a system with an accretion disc, and not as pure stellar emission. The modelling of Popham and Di Stefano (1996) served as further evidence of an accretion disc in CAL 83 (see §2.4.5).

CAL 83 was observed with the Goddard High Resolution Spectrograph (GHRS) on the *Hubble Space Telescope* (*HST*) on 10 November 1996 (Gänsicke et al., 1998,b). These authors compiled a long-term UV lightcurve of CAL 83 by making use of all existing *IUE* observations (spanning 1984-1991) and the *HST* observation, and the bright state of January 1987 clearly stands out (Fig. 3.5).

The *HST* spectrum is shown in Fig. 3.6. The broad Ly α absorption feature is primarily caused by interstellar absorption along the line of sight. Using this line, Gänsicke et al. (1998) determined the total neutral hydrogen density between us and

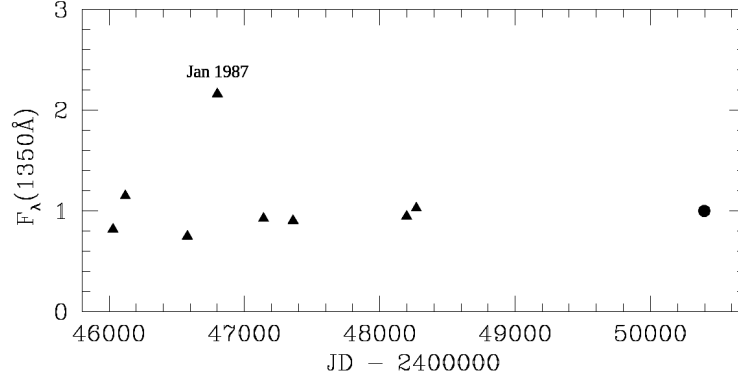


Fig. 3.5: Ultraviolet lightcurve of CAL 83, compiled from *IUE* (triangles) and *HST* (filled circle) data. The values on the vertical axis represent the average flux in the 1300-1400 Å range in units of $10^{-14} \text{ erg cm}^{-2} \text{ s}^{-1} \text{ Å}^{-1}$. (Adopted from Gänsicke et al., 1998, Fig. 2.)

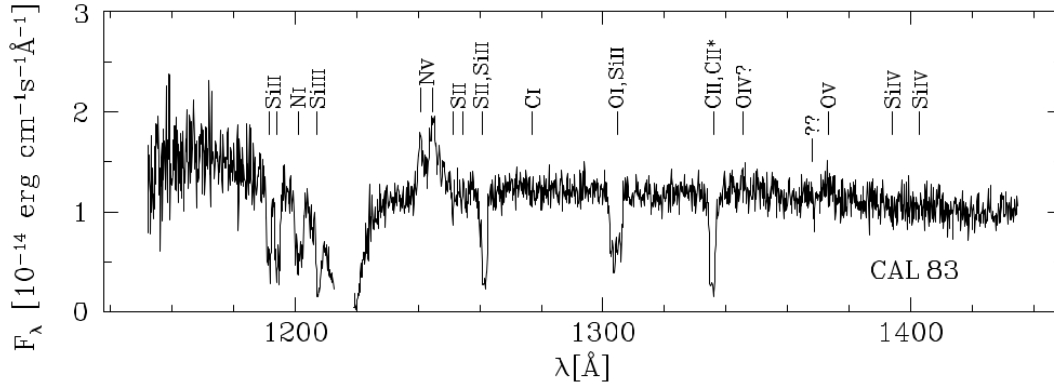


Fig. 3.6: *Hubble Space Telescope* UV spectrum of CAL 83 on 10 November 1996. The geocoronal emission line has been truncated. (Adopted from Gänsicke et al., 1998, Fig. 1.)

CAL 83 to be $N_{\text{H}} = (6.5 \pm 1.0) \times 10^{20} \text{ cm}^{-2}$. This includes interstellar absorption in the Galaxy and the LMC, and also any intrinsic absorption taking place in the binary system itself. As discussed in the previous section, these authors also obtained a best fit column density towards CAL 83 of $7.0 \times 10^{20} \text{ cm}^{-2}$ with their LTE WD atmosphere model fits to the *ROSAT* data, confirming the accuracy of the N_{H} value.

These values are also in excellent agreement with those determined for the galactic foreground column density computed with EXSAS by Zimmermann et al. (1994) from the data of Dickey and Lockman (1990) ($6.5 \times 10^{20} \text{ cm}^{-2}$), as well as the estimates for galactic foreground towards the LMC based on H I 21 cm observations ($\sim 3\text{--}8 \times 10^{20} \text{ cm}^{-2}$) (Fitzpatrick, 1986; Bessell, 1991). From the data of Rohlfs et al. (1984), the total neutral hydrogen column density *within* the LMC along the line of sight to CAL 83 is $\sim 15 \times 10^{20} \text{ cm}^{-2}$. One can conclude that CAL 83 is located in the near side of the LMC, and that there is not substantial intrinsic neutral hydrogen absorption in the source.

The absorption features detected in the *IUE* spectra are also visible in the *HST* spectrum, together with additional strong absorption lines in the previously unexplored region on the far blue side of the spectrum. All these absorption features can be ascribed to interstellar absorption in the direction of the LMC. The most prominent emission features in the spectrum are the lines of the N v $\lambda\lambda 1239, 1243$ resonance doublet. These lines have a FWHM of ~ 1 Å, with wing components extending towards the red up to $\sim +800$ km s $^{-1}$. These lines are expected to originate from the accretion disc, and their small widths indicate that the orbital inclination of CAL 83 is relatively low. The red wing structures may have the same outflow-related origin as the blue- or redshifted wing of the He II $\lambda 4686$ emission.

The emission feature at ~ 1666 Å is present in the *HST* spectrum as well, but its origin is still unclear. If it represents a redshifted component of the He II $\lambda 1640$ line, it indicates a velocity of approximately $+4000$ km s $^{-1}$.

Schmidtke et al. (2004) reported on 11 observations of CAL 83 obtained with the Far Ultraviolet Spectroscopic Explorer (*FUSE*). The only line in the spectrum is the O VI $\lambda 1032$ emission line. The radial velocity semi-amplitude of this line is 23 ± 6 km s $^{-1}$ around an average of $+434$ km s $^{-1}$ (see Fig. 3.16 in the next section). This average is significantly higher than the systemic velocity of the LMC, and it was suggested that this velocity offset may be caused by the blue edge of the line undergoing P Cyg-like absorption. Schmidtke et al. also found that the UV continuum flux varies with the orbital period (again, see Fig. 3.16).

3.1.3 Optical

The central source

A blue object with $B \sim 16.8$ was identified as the optical counterpart of CAL 83 by Cowley et al. (1984). The ICRS coordinates of the source are 05:43:34.5 – 68:22:18.

Pakull et al. (1985) obtained optical spectra of CAL 83 at the ESO 3.6-m telescope at La Silla (Fig. 3.7). They found emission lines typical of an accretion disc, with He II and Balmer lines dominating the spectrum. They pointed out the superficial resemblance between the optical spectrum of CAL 83 and that of the LMXBs LMC X-2 and Sco X-1, and further also remarked on the similarity between the soft X-ray spectra of CAL 83 and that of the black hole candidates LMC X-1 and X-3. They performed various spectroscopic measurements of the source during 1980, 1983 and 1984, and they found that the ratio of He II $\lambda 4686$ to H β emission is always larger than 5, and that the equivalent width of the H β line does not become stronger than -6 Å. In general, LMXBs also exhibit high He II $\lambda 4686$ to H β emission ratios as compared to cataclysmic variables (CVs) (van Paradijs and Verbunt, 1984).

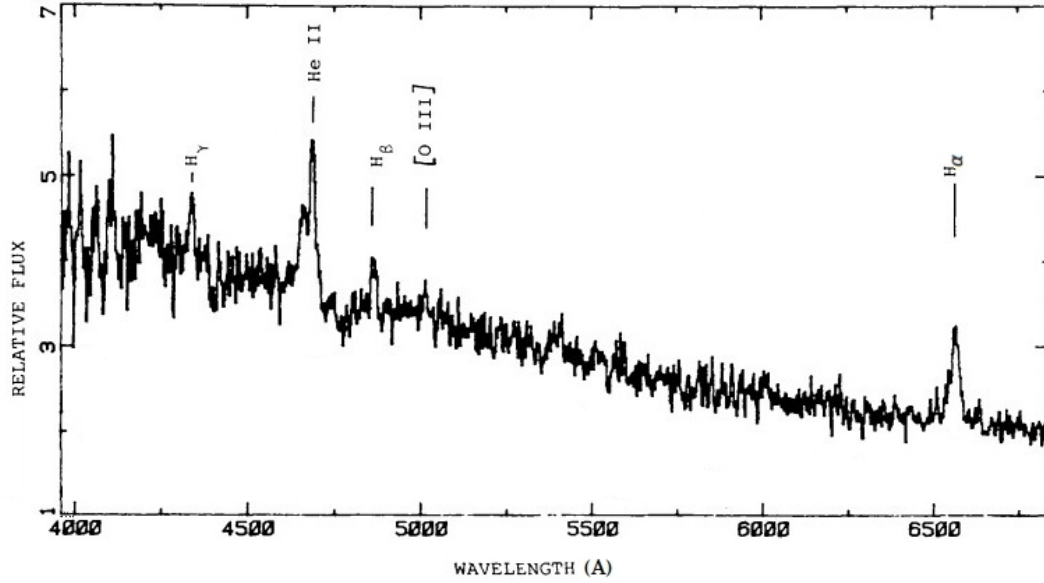


Fig. 3.7: ESO 3.6-m spectrum of CAL 83, obtained in May 1984 with a resolution of ~ 9 - 12 Å. (Adopted from Pakull et al., 1985, Fig. 1.)

Pakull et al. (1985) also reported a blueshifted emission feature around 4660 Å, which could not be completely explained by the O II, C III and C IV multiplets often found at this position, because the associated transitions at other wavelengths were not present with the expected intensity. As a possible explanation of this blueshifted feature, they suggested either the quadratic Zeeman effect due to a magnetic field in the emission region of $\sim 10^7$ G, or a Doppler shift due to velocities of 1500 - 2300 km s $^{-1}$ towards the observer, indicating considerable mass loss in an outflow. They also remarked on the forbidden [O III] emission, which is expected to originate in the ejected material that has been excited by the soft X-ray radiation from the central source.

They performed photometry on CAL 83 during 1980 and 1982 with the 0.9-m Dutch, 1.5-m and 3.6-m ESO telescopes, obtaining average colour indices of $B - V = 0.02 \pm 0.06$ and $U - B = -1.04 \pm 0.06$, which are reconcilable with a blackbody temperature of $20\,000$ K, and similar to the indices measured for LMXBs. They also found erratic brightness changes of 0.3 mag in less than an hour, with a total range of 16.9 - 17.5 in the B -band, but no apparent periodicities. They calculated an absolute visual magnitude of $M_V = -1.2$.

Crampton et al. (1985) obtained spectra of CAL 83 with the Las Campanas 2.5-m du Pont telescope on ten nights from 27 January to 5 February 1984. (Fig. 3.8). Again, the strongest line is the He II $\lambda 4686$ emission line, with a mean equivalent width of -5 Å, and they estimated a radial velocity semi-amplitude of ~ 35 km s $^{-1}$

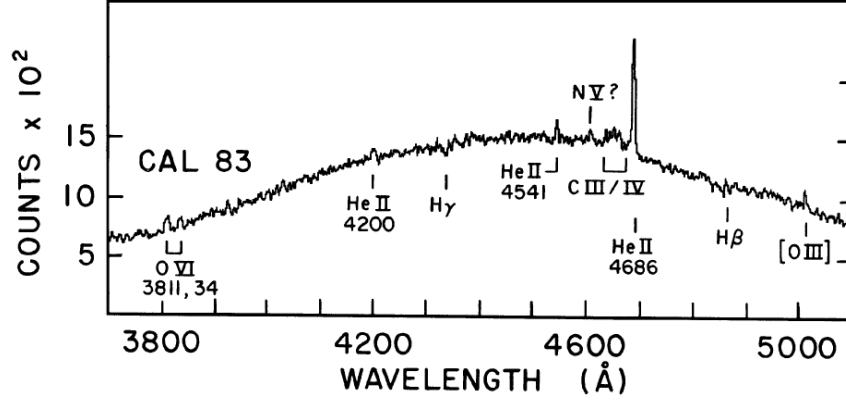


Fig. 3.8: Average spectrum of CAL 83 obtained at the Las Campanas 2.5-m du Pont telescope during 27 January - 5 February 1984 with a resolution of 1.8 Å. (Adopted from Crampton et al., 1985, Fig. 2.)

Table 3.2: Average equivalent widths of CAL 83 emission lines. Note that the values should be interpreted as being negative, as these are emission features. (Adopted from Crampton et al., 1987, Table 2.)

Epoch	H δ	H γ	H β	H α	C III/IV	He II	O VI
1982 Feb	...	...	...	...	...	10.9	...
1982 Nov	...	...	2.4	...	...	10.0	1.2
1984 Jan	...	...	...	...	2.5	5.0	1.8
1985 Mar	2.6	2.3	6.4	...	2.8	12.8	0.7
1985 Dec	1.5	1.8	5.3	18	2.2	12.2	...

for this line.

A detailed spectroscopic study of CAL 83 at the Cerro Tololo Inter-American Observatory (CTIO) and Las Campanas Observatory was undertaken by Crampton et al. (1987) over a period of 4 years (1982-1985) (see Fig. 3.9). Portions of their average spectra in the blue region for five observing runs are shown in Fig. 3.10, and the measured equivalent widths are listed in Table 3.2. These authors report rapid variations in the main He II λ 4686 emission line on time-scales of hundreds of seconds, which was also observed by Charles, Smale and van Paradijs (1983, private communication with Crampton et al., 1987).

Many of their spectra contained the same blue wing structure in He II λ 4686 that was observed by Pakull et al. (1985). However, it appeared that this wing is variable; in fact it shifted towards the red side of the main emission line during the 1985 observations, and a similar wing was seen in H β .

When the wing structure of He II λ 4686 was redshifted, it revealed weak emission features remaining at 4660 Å. From the Crampton et al. (1987) spectra, it therefore appeared that the blue wing of He II λ 4686 consists of both a Doppler shifted com-

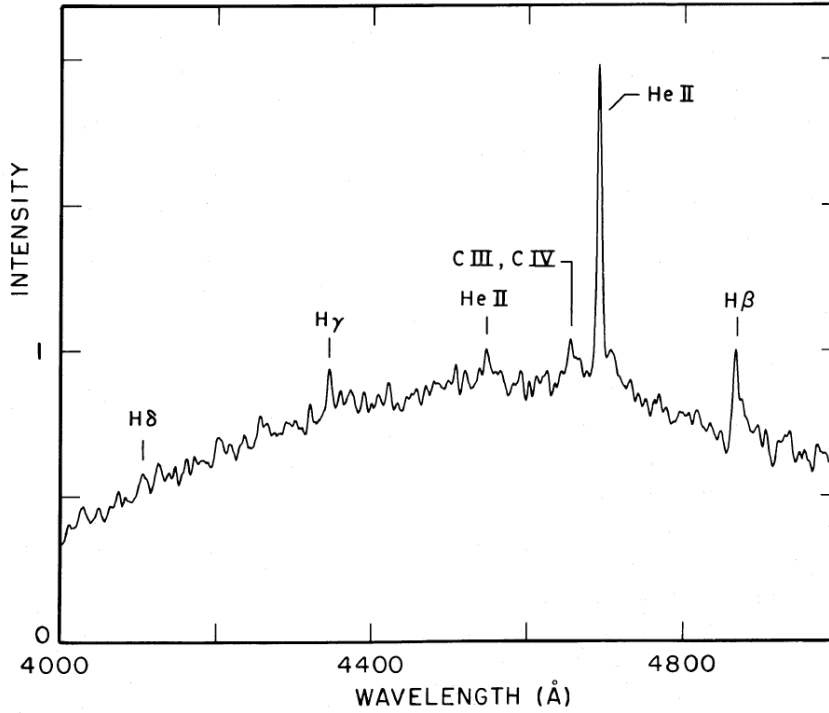


Fig. 3.9: Average spectrum of CAL 83 obtained during December 1985 with the “2D-Frutti” detector at Las Campanas. The resolution is ~ 2.5 Å. (Adopted from Crampton et al., 1987, Fig. 1.)

ponent from some kind of outflow, and weak emission features from ionized carbon (C III and C IV) and probably also nitrogen as part of the Bowen blend (a blend of N III $\lambda\lambda 4634, 4640$ and C III $\lambda\lambda 4647, 4650$, e.g. Cornelisse et al., 2013). The presence of highly ionized carbon is supported by the presence of O VI emission lines, and indicate emission from a very hot star or accretion disc. Crampton et al. (1987) also pointed out the low ratio of hydrogen emission to He II and CNO emission. Within the accreting binary model, this serves as evidence of CNO cycling in the donor star, indicating an evolved donor of which the outer layers have already been stripped by accretion.

The broad He II $\lambda 4686$ wing structure did not appear to be variable on time-scales of up to 10 days, but rather on a longer time-scale. Crampton et al. (1987) suggested that this variable structure may be associated with an outflow (either a wind or a collimated outflow) from a precessing accretion disc. In such a case, the precession period may be ~ 69 d, as tentatively derived from the variation in the radial velocity of the He II wing. The broadness of the wing indicates that, if it is associated with a collimated outflow, the opening angle has to be quite large, or it should contain a large range of velocities.

Southwell et al. (1997) compared the known characteristics of CAL 83 with the con-

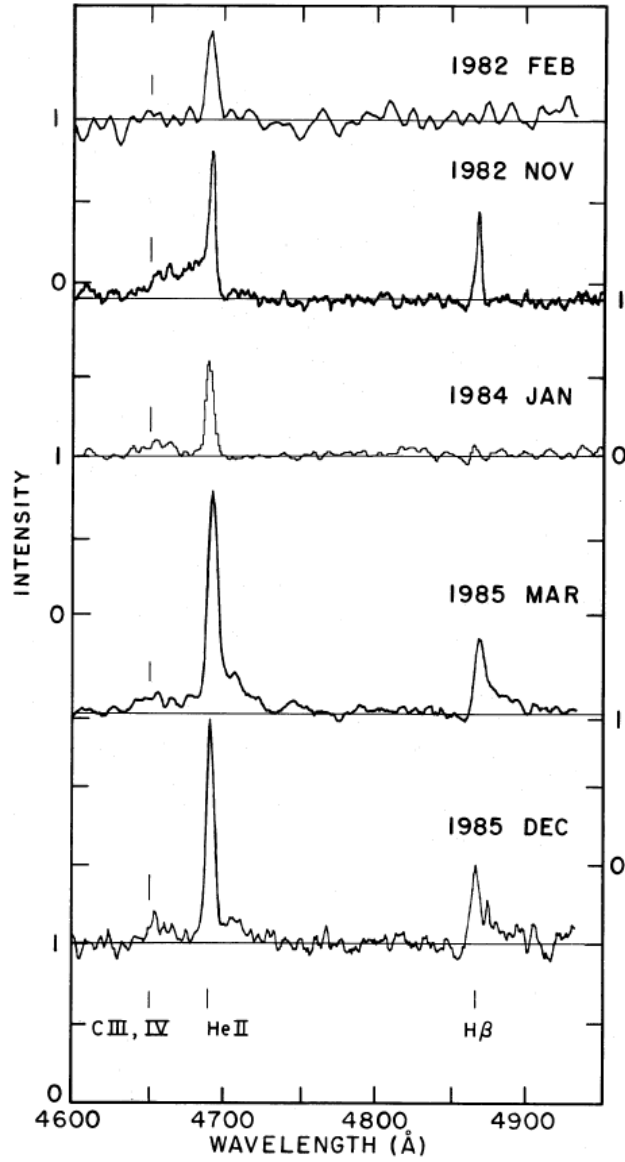


Fig. 3.10: Average spectra of CAL 83 from 5 observing runs at the CTIO and Las Campanas. Note the changing wing structure of the He II $\lambda 4686$ and H β emission lines. (Adopted from Crampton et al., 1987, Fig. 2.)

ditions for irradiation-induced disc warping discussed in §2.4.4, and found that the accretion disc in CAL 83 should be unstable against disc warping, and that the expected time scale of the precession is of the order of a few months. Although the precession period of ~ 69 d mentioned above has not been confirmed, it is in good agreement with the theoretical prediction.

From the radial velocity curve of He II $\lambda 4686$, Crampton et al. (1987) derived an orbital period of 0.93036 ± 0.00006 d with an associated semi-amplitude of $K_1 = 31 \pm 5$ km s $^{-1}$, while not ruling out a possible orbital period of 1.035 d. The systemic velocity is 270 ± 4 km s $^{-1}$, which is reconcilable with its LMC membership. They also used K_1 to calculate the mass function (see §2.2) for various possible inclinations, and they found that for any mass of the primary star, the secondary star will be less massive. They also estimated the inclination to be between 20° and 30° .

However, it is important to remember that measurements of K_1 from the radial velocity modulation of the He II $\lambda 4686$ emission line assume that this line provides an exact measure of the primary orbital motion, which would only be true if the distribution of He II emission was perfectly symmetrical about the primary star. There are, unfortunately, serious problems with this approximation, as discussed by e.g. Roelofs et al. (2006); Cantrell and Bailyn (2007), compromising the accuracy of such a mass function calculation. As pointed out by Cowley et al. (1998), the high-excitation O VI emission lines weaken or disappear near minimum light in the eclipsing sources, and are therefore expected to originate very close the primary. In principle, this would make them better candidates for obtaining the radial velocity curve of the primary, but unfortunately these lines are very faint, so in practice this is very difficult to accomplish.

Photometry with the CTIO 1.5-m telescope during 11-14 November 1985 yielded, on average, $V = 17.28 \pm 0.03$ and $B - V = -0.025 \pm 0.034$ (Crampton et al., 1987). Some variability was apparent in both the V and B filters, and in particular on 12 November, a series of B -magnitude measurements revealed a smooth rise and decline with an amplitude of 0.15 mag and a duration of 1.2 h.

Crampton et al. (1987) briefly considered the possibility of CAL 83 consisting of a WD in a CV system. However, the X-ray luminosity would then put it in the foreground in the Galactic halo, which would mean the large Roche-lobe filling secondary should be clearly visible in the red part of the spectrum, which it is not. Therefore they did not favour this theory, but rather considered CAL 83 as an unusual LMXB, with the low X-ray to optical flux possibly due to the secondary not filling its Roche lobe completely, yielding a low mass transfer rate, or due to the

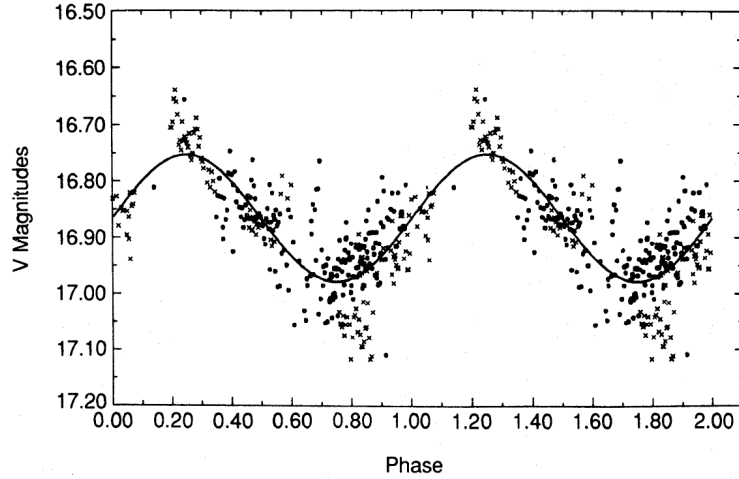


Fig. 3.11: December 1984 lightcurve of CAL 83, folded on $P_{\text{orb}} = 1.0436$ d. Two orbital cycles are shown for clarity. (Adopted from Smale et al., 1988, Fig. 3.)

central X-ray source being occulted by the accretion disc.

After an extensive photometric observing campaign during December 1984, mainly with telescopes at the SAAO and ESO, Smale et al. (1988) unambiguously determined the orbital period of CAL 83 to be 1.0436 ± 0.0044 d, with a semi-amplitude of 0.113 ± 0.012 mag. (Utilizing the spectroscopic data of Crampton et al., 1987, Cowley et al., 1991 later refined this period to $P_{\text{orb}} = 1.04752 \pm 0.00008$ d.) Their folded lightcurve is shown in Fig. 3.11. They also reported variability of up to 0.2 mag on a time-scale of ~ 2 hours in these lightcurves. After comparing the magnitudes of the December 1984 observing run with those of another run in August 1984, Smale et al. (1988) found that CAL 83 was considerably brighter in August (by ~ 0.5 mag). These authors also considered the possibility of CAL 83 being an LMXB, albeit with a much lower X-ray to optical luminosity ratio of $L_X/L_{\text{opt}} \sim 0.7$ as compared to the typical value of $\sim 10^2$ - 10^3 in LMXBs.

Smale et al. (1988) also presented spectroscopy obtained at ESO, the SAAO and with the Anglo-Australian Telescope (AAT) (see Fig. 3.12). The principal feature is the He II $\lambda 4686$ line, with H β also in emission. The region around the He II emission line is shown in Fig. 3.13 for the higher-resolution ESO and AAT spectra, and the varying blue wing structure is evidently also present here. The emission line parameters are provided in Table 3.3, and one can see that there is also significant variability in the widths and intensities of the main He II $\lambda 4686$ and H β lines.

By folding the radial velocity variations of He II $\lambda 4686$ on the photometric period, Smale et al. (1988) found a semi-amplitude of $K_1 \sim 40$ km s $^{-1}$ around the systemic velocity of ~ 290 km s $^{-1}$. Once again, such a low value of K_1 leads to a small mass

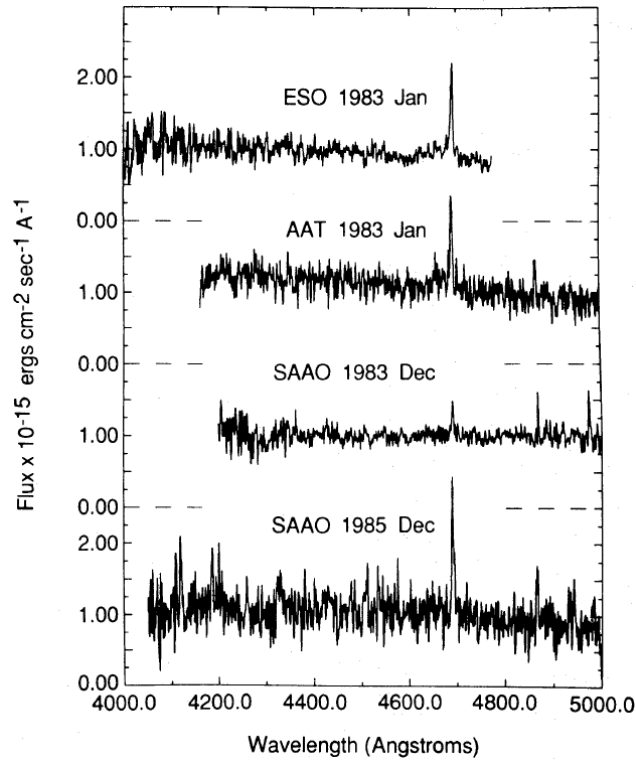


Fig. 3.12: Optical spectra of CAL 83. (Adopted from Smale et al., 1988, Fig. 5.)

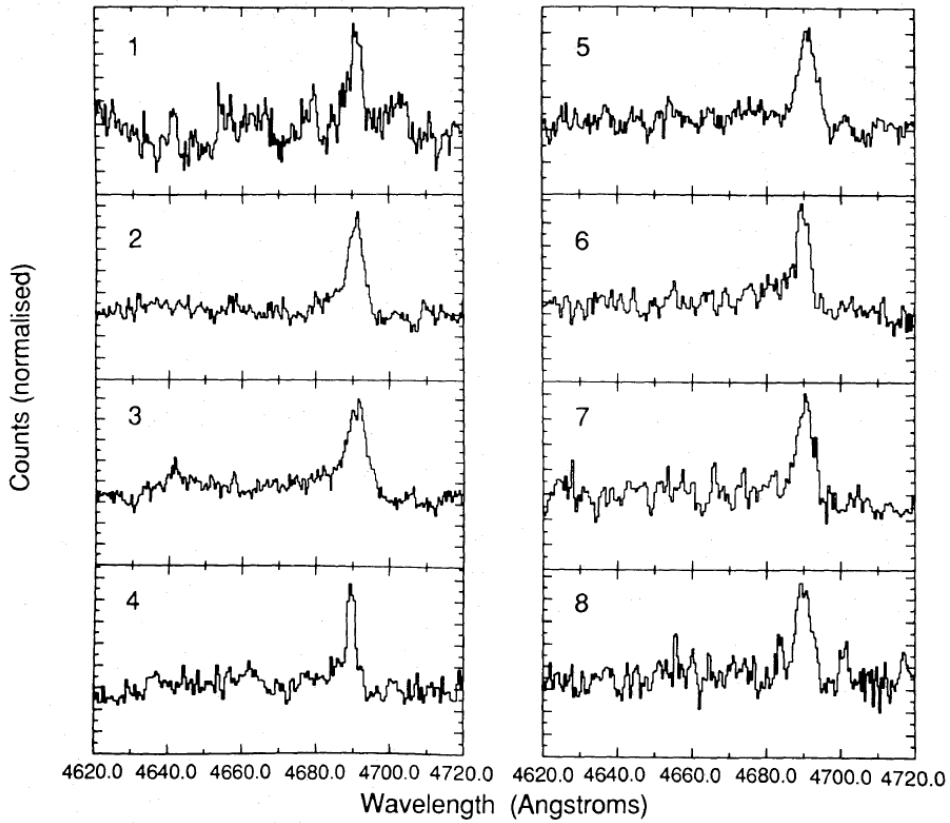


Fig. 3.13: The region around He II $\lambda 4686$ in ESO and AAT spectra of CAL 83 (see Table 3.3 for more information). (Adopted from Smale et al., 1988, Fig. 6.)

Table 3.3: Parameters of the emission lines of CAL 83 in the spectra in Fig. 3.12 and Fig. 3.13. For the He II line, the parameters are for the main emission feature, excluding the wing structure. The equivalent widths should be interpreted as negative values, as these are emission features. (Adopted from Smale et al., 1988, Table 4.)

Date	U T	Obs	He II $\lambda 4686$			H β		
			Velocity (km/sec)	FWHM (km/sec)	Eq.Width (Å)	Velocity (km/sec)	FWHM (km/sec)	Eq.Width (Å)
1983 Jan 20, 04:57	(1)	ESO	325±25	242±28	9.1±2.5			
Jan 20, 05:43	(2)	ESO	334±24	242±24	7.6±1.9			
Jan 21, 05:58	(3)	ESO	341±20	359±22	8.9±1.5			
Jan 21, 16:18	(4)	AAT	252±30	184±29	4.8±0.9			
Jan 22, 03:24	(5)	ESO	336±22	358±23	8.2±1.7			
Jan 22, 14:59	(6)	AAT	259±21	271±20	6.7±1.4			
Jan 23, 12:25	(7)	AAT	298±25	318±29	8.1±1.6			
Feb 1, 14:05	(8)	AAT	273±23	319±26	7.1±1.8			
1983 Jan	(Sum)	ESO	326±16	376±37	8.2±1.1	326±16	376±37	8.2±1.1
1983 Jan/Feb	(Sum)	AAT	260±20	340±33	7.3±0.9	255±19	160±40	1.8±0.5
1983 Dec	(Sum)	SAAO	233±23	279±52	2.4±0.4	402±40	93±10	1.4±0.3
1985 Dec 16	(Sum)	SAAO	235±23	407±52	12.5±2.3	229±34	362±74	4.2±1.7

function, requiring the mass of the secondary to be smaller than the mass of the primary. Specifically, for a neutron star primary, the secondary is required to have a mass smaller than $0.45 M_{\odot}$ and can only fill its Roche lobe if it is an evolved star. The companion can only be a main sequence star if the primary mass is much larger than $3 M_{\odot}$, i.e. if CAL 83 contains a black hole.

Cowley et al. (1998) presented spectra of CAL 83 obtained with the CTIO 4-m telescope during 3 nights in November 1996 (Fig. 3.14). In these spectra, and also in a spectrum reported by Charles, Southwell and Livio (1997), the red wings of the H and He II emission lines are enhanced. They also pointed out the P Cyg profiles in the lines, corresponding to an outflow velocity of approximately -690 km s^{-1} . Together with the previous CAL 83 spectra, these wing structures are still reconcilable with the $\sim 69 \text{ d}$ outflow precession period suggested by Crampton et al. (1987). The spectral line properties are given in Table 3.4.

Cowley et al. (1998) also reported new photometric measurements of CAL 83 from 1993, 1995 and 1996, and remarked on the considerable variation in the optical magnitude. However, at that stage it was not yet possible to determine whether these variations had any regular pattern.

Although Greiner and Di Stefano (2002) found no significant variability in the $V - R$ colour of the MACHO lightcurves, a re-analysis by Schmidtke et al. (2004) showed

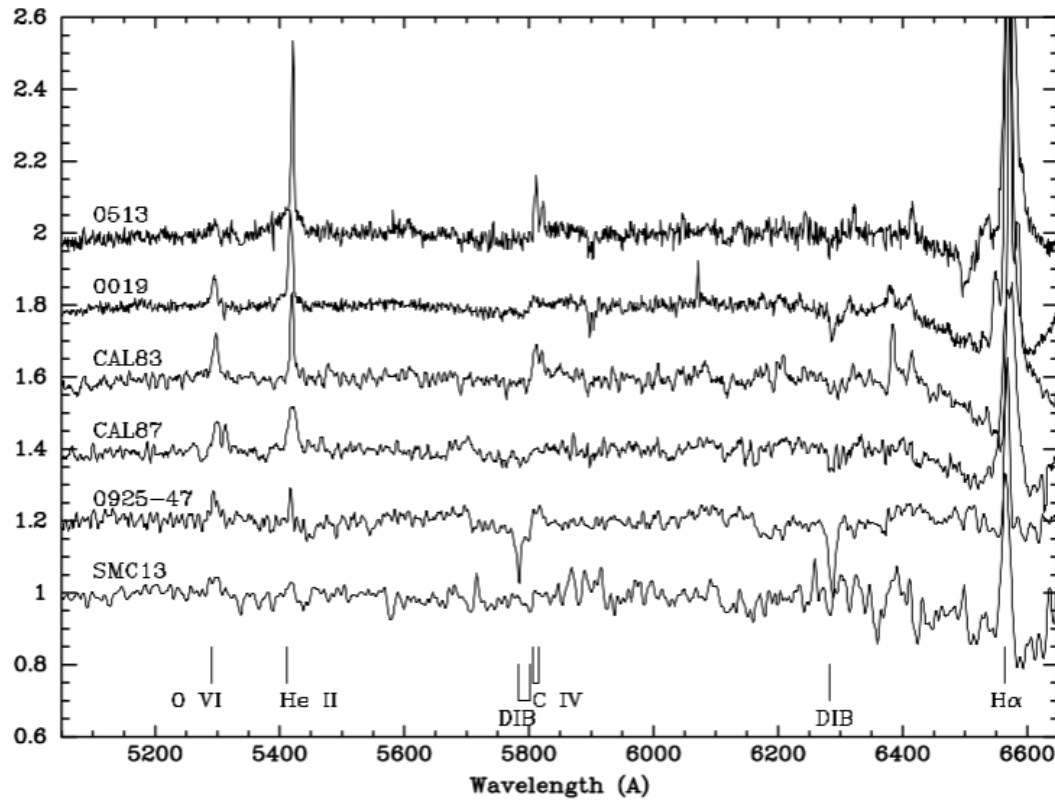
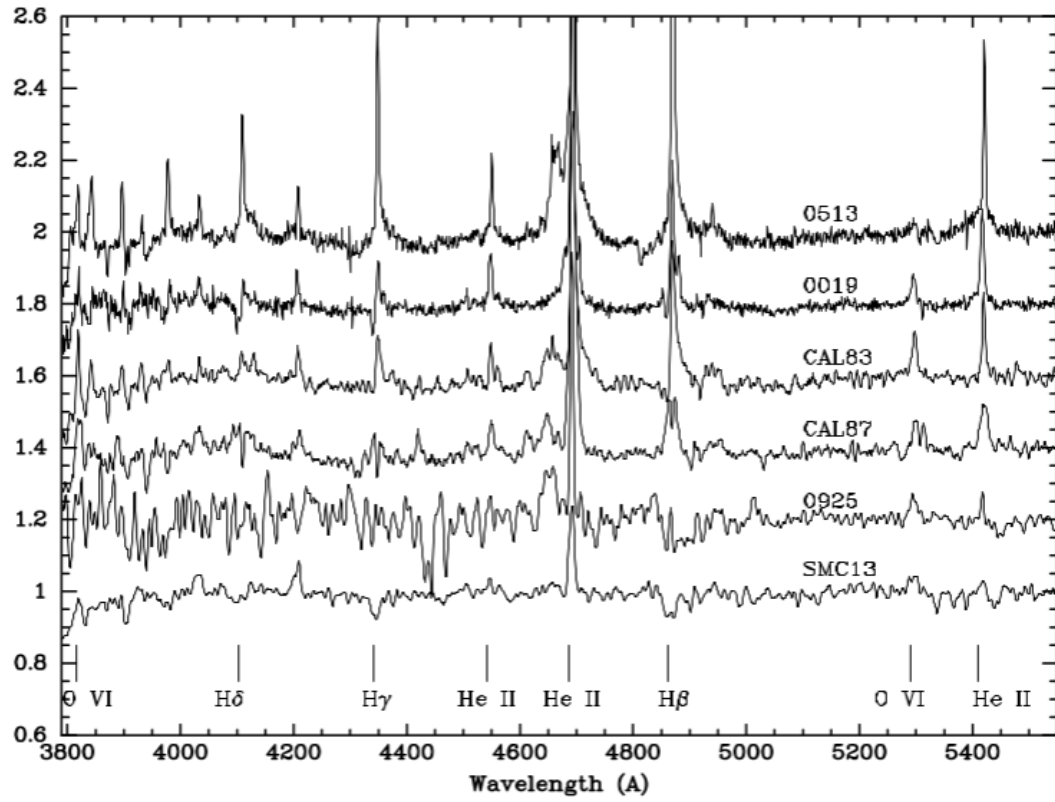


Fig. 3.14: Optical spectra of SSSs obtained with the CTIO 4-m telescope in November 1996. (Adopted from Cowley et al., 1998, Fig. 1.)

Table 3.4: The average equivalent widths of the CAL 83 emission lines as observed with the CTIO 4-m telescope in November 1996 (Cowley et al., 1998, Table 5). The Balmer lines may include contributions from the He II Pickering lines. The weaker O VI λ 3835 emission is not included in the O VI λ 3811 measurement.

Line ID	Measurement ($\text{\AA}$)
H α	−21.9
H β	− 4.8
H γ	− 1.9
He II λ 4686	−12.1
He II λ 4541	− 0.8
He II λ 5411	− 1.8
O VI λ 5290	− 1.3
O VI λ 3811	− 1.6
FWHM of He II λ 4686	5.4

that there is a slight dependence of colour on the optical state, with the average colour in the optical high state being bluer than the average colour in the intermediate state, which is again bluer than the low state. Schmidtke et al. derived a new orbital ephemeris from the MACHO lightcurves, i.e.

$$T_{\min} = \text{JD } 2\,451\,500.953 \pm 0.004 + (1.047568 \pm 0.000003)E \text{ days}, \quad (3.1)$$

where $T_{\min}$ is the time of minimum light and E is an integer. This is currently the most up-to-date complete orbital ephemeris of CAL 83. New photometric measurements in the B and V bands obtained by Schmidtke et al. (2004) with the CTIO 0.9-m telescope in March 1999 are in good agreement with this ephemeris. It was also shown that the semi-amplitude of the orbital modulation increases when the overall optical brightness of the system is higher (see Fig. 3.15).

With their new orbital ephemeris, Schmidtke et al. (2004) performed a re-analysis of the optical spectra of Crampton et al. (1987), and obtained the revised orbital elements listed in Table 3.5. In Fig. 3.16, the *FUSE* continuum level is shown to vary with the MACHO orbital period, and the radial velocity curves of O VI λ 1032 and He II λ 4686 are compared. It is obvious that the relative phasing of these lines are not the same, therefore they probably originate in different parts of the system. They also do not have the same phasing as the photometric lightcurve.

Oliveira and Steiner (2006) reported the detection of the [Fe x] λ 6375 emission line in CAL 83, the first time that this line was seen in a close binary supersoft source. They confirmed the presence of outflow signatures in the H and He II lines as seen before, and also found an emission feature at 6548 $\text{\AA}$ that may be [N II], or perhaps a blueshifted component of H α .

CAL 83 was monitored with the 1.3-m Warsaw telescope at Las Campanas Obser-

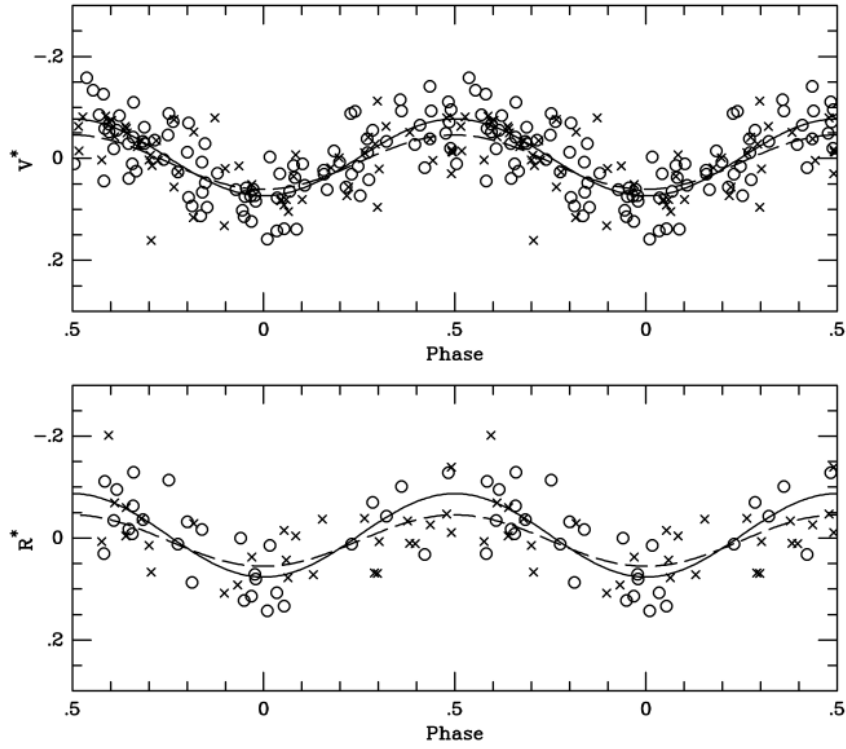


Fig. 3.15: MACHO lightcurves of CAL 83, folded on an orbital period of 1.047568 d. Crosses indicate data points from a MACHO data segment between MJD 48917 and 49 640 (intermediate state), and circles data from the segment between MJD 50980 and 51541 (high state). The fitted sine waves indicate the semi-amplitude of the system in its lower (dashed line) and higher (solid line) states respectively. (Adopted from Schmidtke et al., 2004, Fig. 4.)

Table 3.5: The revised orbital parameters for CAL 83, based on He II $\lambda 4686$ observations by Crampton et al. (1987). (Adopted from Schmidtke et al., 2004, Table 2.)

Parameter	Revised Value
Time of maximum velocity	JD 2,446,412.785 $\pm$ 0.008
Photometric phase of max. vel.	$\Phi = 0.875 \pm 0.008$
Velocity semi-amplitude	$K = 27.0 \pm 1.6 \text{ km s}^{-1}$
Systemic velocity	$V_0 = +257.0 \pm 1.0 \text{ km s}^{-1}$
Orbital period	$P = 1.047568 \pm 0.000003 \text{ days (fixed)}$
Orbital eccentricity	$e = 0 \text{ (assumed)}$

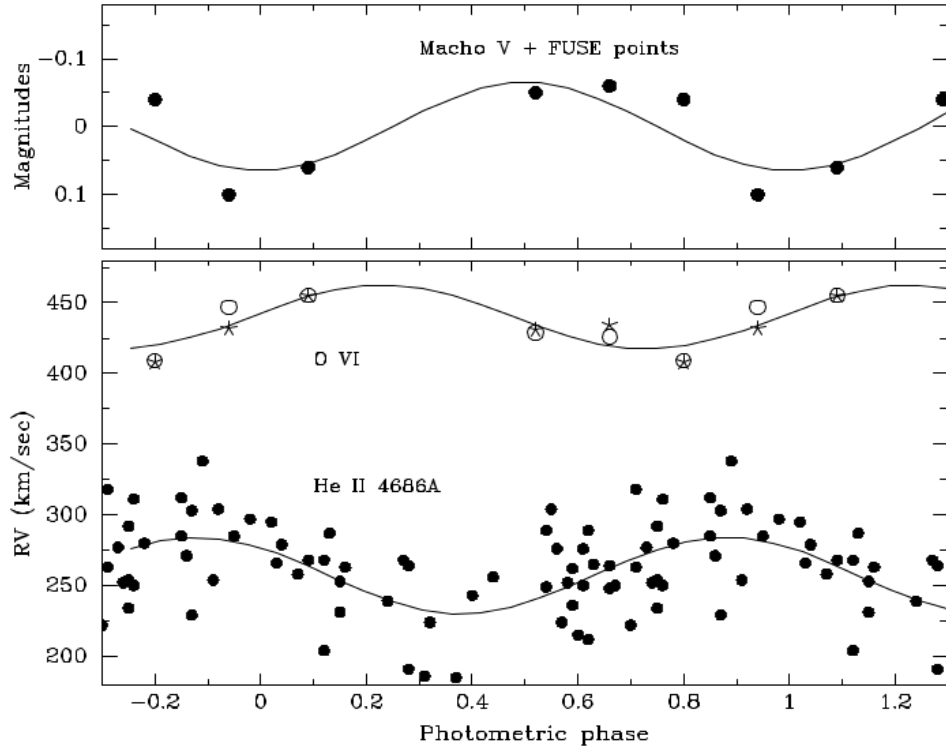


Fig. 3.16: The *FUSE* UV continuum flux, as well as the radial velocity (RV) curves of O VI $\lambda 1032$ and He II $\lambda 4686$, plotted versus the MACHO orbital phase. (Adopted from Schmidtke et al., 2004, Fig. 6.)

vatory as part of the OGLE project (phase III) from 2001 to 2009. Rajoelimanana et al. (2013) investigated the long-term variability of the source by analysing the MACHO and OGLE-III lightcurves, which are shown in Fig. 3.17. By performing a Lomb-Scargle period analysis of these lightcurves, they derived a quasi-period of 450 ± 3 d for the long-term variability (Fig. 3.18). The durations of the optical low and high states are ~ 200 and ~ 250 d, with absolute visual magnitudes M_V of approximately -1 and -2 respectively. The transitions between states typically last ~ 5 -10 d. Rajoelimanana et al. also derived an orbital period of 1.047529 ± 0.000001 d from the OGLE-III lightcurve (Fig. 3.19).

The surrounding ionized nebula

From a theoretical investigation, Rappaport, Chiang, Kallman and Malina (1994) predicted that SSSs should be surrounded by ionized regions, with the central X-ray emitter as the ionizing source. These ionization nebulae are expected to have properties quite different from other astrophysical gaseous nebulae. In particular, they should be very bright in [O III] $\lambda 5007$, with typical intensity ratios of [O III] $\lambda 5007/\text{H}\beta \sim 12$, He II $\lambda 4686/\text{H}\beta \sim 0.47$, [O I] $\lambda 6302/\text{H}\beta \sim 5$ and [N II] $\lambda 6585/\text{H}\beta \sim 6.4$ for a solar composition. And indeed, Pakull and Motch (1989) presented a [O III] $\lambda 5007$ CCD image of CAL 83, showing the highly ionized,

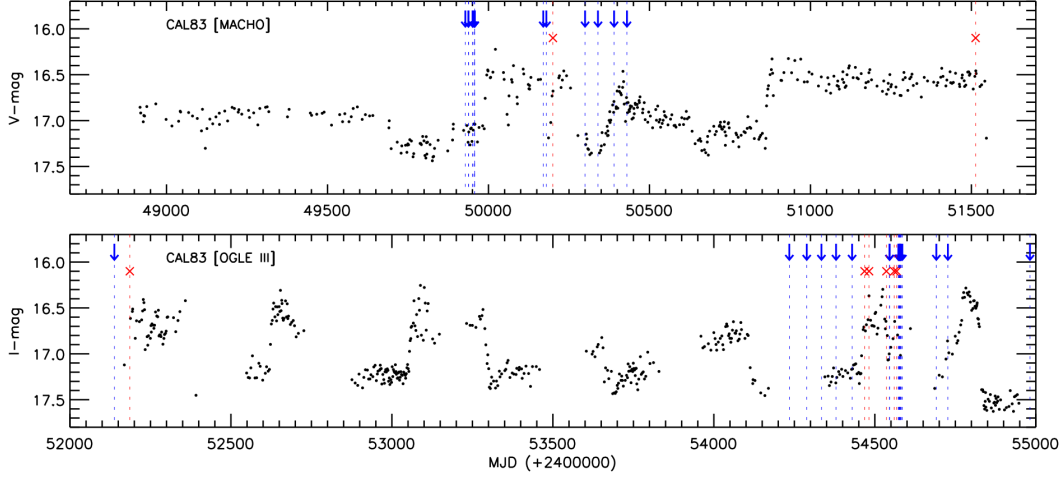


Fig. 3.17: The MACHO and OGLE-III lightcurves of CAL 83. The quasi-periodic cycles recur every ~ 450 d, with the optical high states lasting ~ 250 d. The blue arrows and red crosses represent times of observed X-ray on- and off-states respectively. (Adopted from Rajoelimanana et al., 2013, Fig. 1)

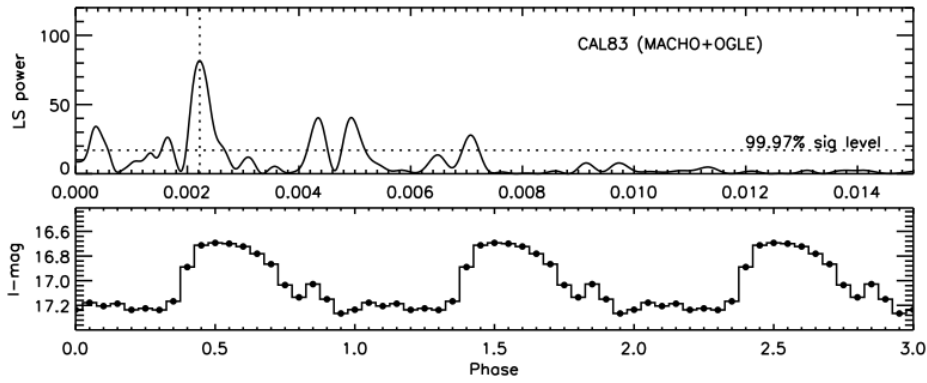


Fig. 3.18: The Lomb-Scargle (LS) periodogram (top) and folded MACHO + OGLE-III lightcurve (bottom) of CAL 83, illustrating the long-term ~ 450 d period. (Adopted from Rajoelimanana et al., 2013, Fig. 3)

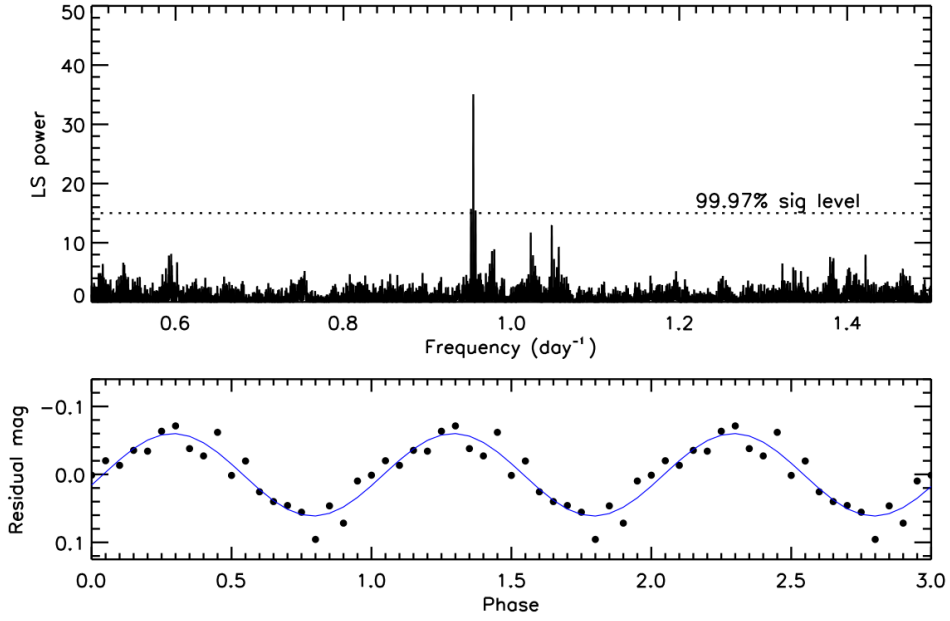


Fig. 3.19: The Lomb-Scargle (LS) periodogram (top) and folded OGLE-III lightcurve (bottom) of CAL 83, illustrating the derived orbital period of 1.047529 ± 0.000001 d. The phasing is arbitrary. (Adopted from Rajoelimanana et al., 2013, Fig. 4)

asymmetric nebular region around the central source, with a diameter of 50 arcsec.

Crampton et al. (1987) confirmed the presence of the [O III] nebular lines at 4959 and 5007 Å in spectra where the emission from the surrounding region has not been subtracted. Smale et al. (1988) also found [O III] $\lambda 4959$ and $\lambda 5007$ emission distributed uniformly along the 1 arcmin slit that they used, confirming the presumed emission from an ionized nebula around CAL 83. However, they did not find any indication of He II or $H\beta$ emission extended beyond the central source. Cowley et al. (1998), however, reported the presence of extended $H\beta$ nebular emission, as well as nebular emission lines of [O II] $\lambda 3727$, $H\gamma$, [O III] $\lambda\lambda 4959, 5007$ and $H\alpha$.

Remillard, Rappaport and Macri (1995) undertook a detailed investigation at the CTIO of the emission lines in the ionization nebula around CAL 83, and their findings were consistent with the theoretical predictions. Their images of the CAL 83 nebula in [O III] $\lambda 5007$ and $H\alpha$ are shown in Fig. 3.20. The nebula has an asymmetric shape, with a bright inner region with a typical radius of ~ 7.5 pc and a density of $\sim 5 \text{ cm}^{-3}$, and a more diffuse outer region extending to ~ 20 pc. These authors estimate that the inner nebula contains $\sim 150 M_{\odot}$, while the material ejected from the binary system itself (e.g. by an accretion wind) should only be of the order of $\sim 1 M_{\odot}$. Therefore the nebula is primarily caused by the ionization of the local ISM by the supersoft X-rays from CAL 83.

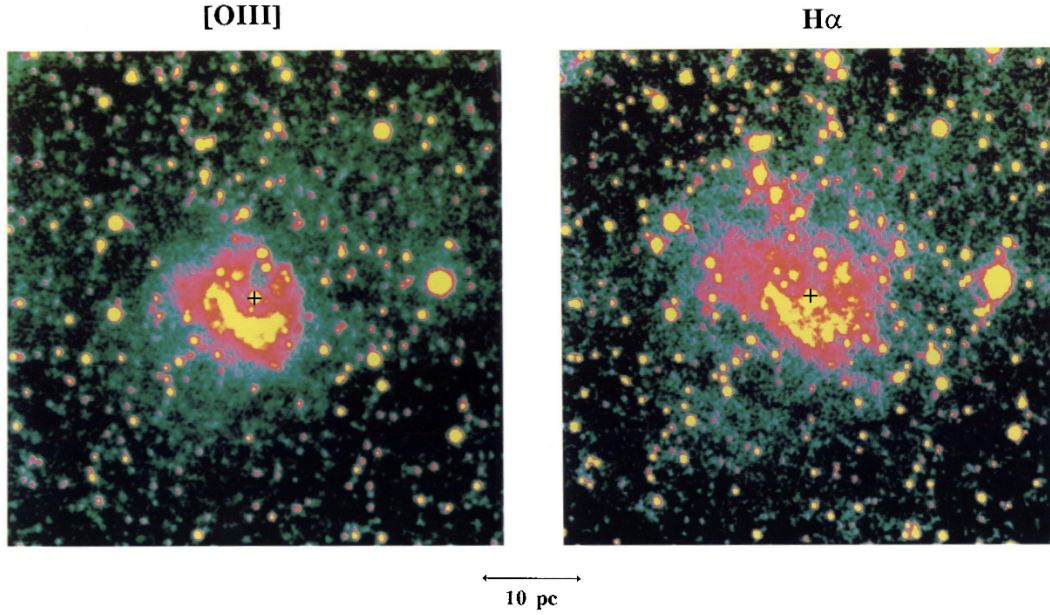


Fig. 3.20: The CAL 83 ionization nebula in the emission lines of $[\text{O III}] \lambda 5007$ (left) and $\text{H}\alpha$ (right). CAL 83 is the central object marked with a cross in each of the images. The artificial colours of black, green, red and yellow represent surface brightness contours with relative scale factors of approximately 1, 3 and 5. With this scale, the sky background is zero and the brightest nebular regions ~ 9 . Each image has a width of 3.1 arcmin, corresponding to 50 pc. North is to the top, east to the left. (Adopted from Remillard et al., 1995, Fig. 1.)

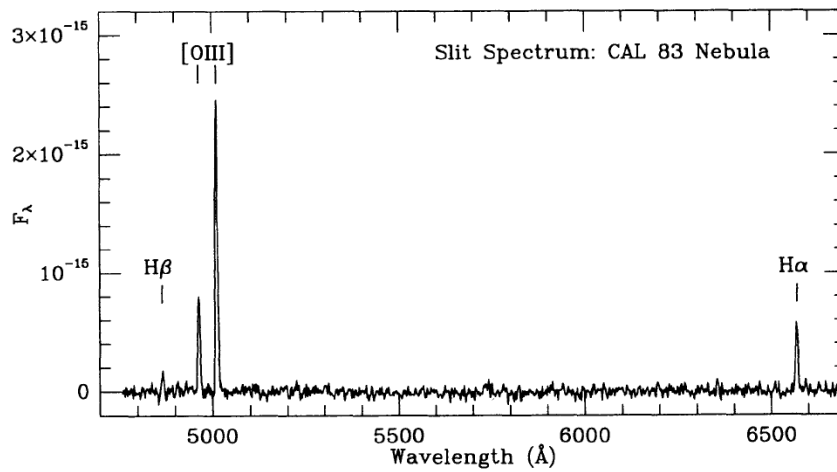


Fig. 3.21: The optical spectrum of the inner nebular region of CAL 83, obtained with a 4 arcsec slit in an east-west orientation. (Adopted from Remillard et al., 1995, Fig. 2.)

Chiang and Rappaport (1996) performed theoretical calculations of the time-dependent properties of ionization nebulae around SSS, considering the ionization evolution of hydrogen and helium, and assuming the density of the gas is uniform at all times. If one assumes that the nebula surrounding CAL 83 can be considered as a fully developed nebula in ionization equilibrium around a “relatively steady” luminous source, and assuming a nebular density of 5 cm^{-3} , their calculations indicate that CAL 83 has been radiating at a time-averaged luminosity of $\gtrsim 3 \times 10^{37} \text{ erg s}^{-1}$ for at least the past $\sim 10^5$ years. Such a high WD luminosity can not be achieved through accretion alone, and this indicates that the source has already passed through an extended period of surface nuclear burning. “Relatively steady” means either a constant luminosity, or a luminosity that is variable on time-scales of $\lesssim 2 \times 10^3 \text{ yr}$.

Using the Very Large Telescope array (VLT) at ESO, Gruyters et al. (2012) performed an in-depth study of the CAL 83 nebula, including the creation of flux maps of the [O III] $\lambda 5007$, [O I] $\lambda 6300$, $\text{H}\alpha$ and [S II] $\lambda 6716$ emission lines, flux ratio maps, and modelling nebulae around SSSs with the CLOUDY ionization code (Ferland et al., 1998). They found that the low ionization ions peak at the edge of the nebula, while the ionization states are higher towards the centre, as expected. They estimated an average electron density of $\sim 10 \text{ cm}^{-3}$ in the nebula, which is similar to the value of Remillard et al. (1995). Gruyters et al. (2012) also reported the first discovery of He II $\lambda 4686$ emission from the nebula itself, with a strange asymmetric distribution around the central source. Their spectrum of the nebula is shown in Fig. 3.22, while the spectrum of CAL 83 without the nebula is shown in Fig. 3.23.

3.1.4 Longer wavelengths

Cowley et al. (1991) undertook a campaign to investigate the possible presence of the Ca II infrared triplet lines at 8498, 8542 and $8662 \text{ \AA}$ in several LMXBs and CVs, with the purpose of measuring the radial velocity curve of the secondary star. For CAL 83, they reported only a marginal detection of Ca II, as well as weak H emission in the near-infrared.

Fender, Southwell and Tzioumis (1998) performed radio observations of X-ray sources in the Magellanic Clouds with the Australian Telescope Compact Array (ATCA) during June-July 1997. Neither CAL 83 nor the nebula surrounding it was detected, and an upper limit of $< 0.12 \text{ mJy}$ was determined for the source at wavelengths of 3.5 cm and 6.3 cm.

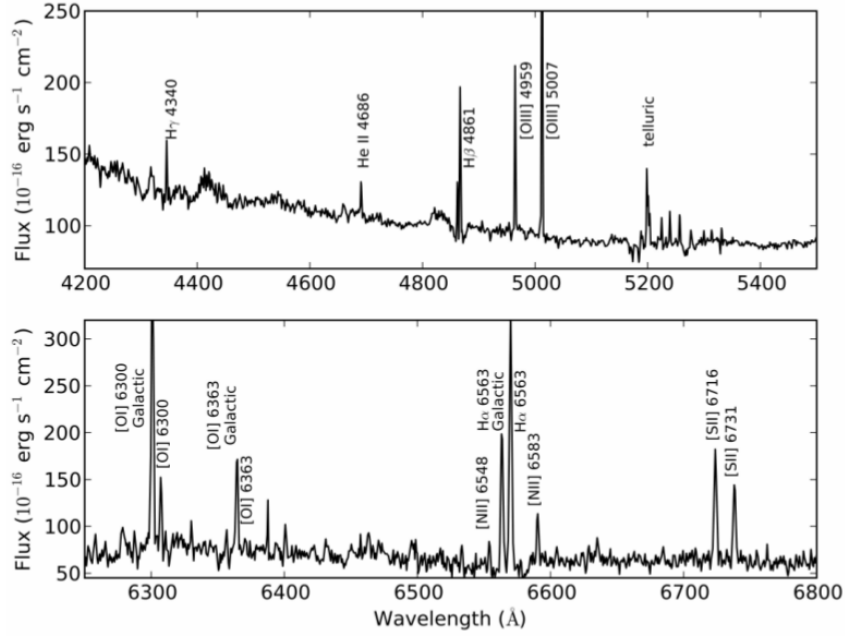


Fig. 3.22: Spectra of the CAL 83 nebula, excluding the central source and surrounding stars, taken from an area of 7.5×7.5 pc² around the source. (Adopted from Gruyters et al., 2012, Fig. 2.)

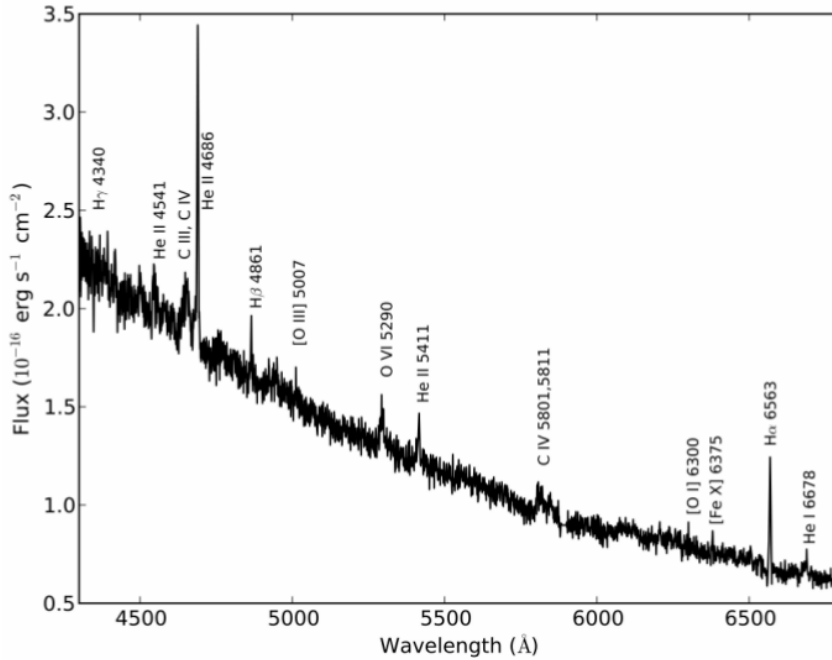


Fig. 3.23: The stellar spectrum of CAL 83, without the contribution from the ISM. (Adopted from Gruyters et al., 2012, Fig. 6.)

3.2 RX J0513.9-6951

3.2.1 X-ray

Schaeidt, Hasinger and Trümper (1993) reported the discovery of RX J0513.9-6951 (RXJ0513) in the *ROSAT* all-sky survey data. This source (often referred to as “Schaeidt’s source”), was observed for 22 days starting on 20 October 1990. The count rate increased from ~ 0.03 to ~ 1 counts s^{-1} during the first few days, whereafter it remained high, but exhibited oscillations with a time-scale of several days for the remaining ~ 10 days’ observations. It was not detected during the *ROSAT* calibration phase of 16-25 June 1990, with a 2σ upper limit of $< 2.23 \times 10^{-3}$ counts s^{-1} for the *ROSAT* 0.07-2.4 keV band. The source was also not detected by the *Einstein Observatory* (Wang et al., 1991).

The calculations of Kato (1985) indicated that considerable expansion and cooling of the envelope of an accreting, nuclear burning WD occurs when the Eddington limit is approached: for a $1.3 M_{\odot}$ WD an increase in accretion rate from 90% to 97% of the Eddington rate will increase the radius and temperature by factors of 4 and 0.5 respectively. If the X-ray flux changes were entirely due to a decreasing WD photospheric radius and not due to a change in luminosity, Pakull et al. (1993) remarked that the non-detection in the *ROSAT* soft X-ray band in June 1990 could be explained if $kT_{\text{eff}} \leq 20$ eV. Pakull et al. therefore explained the X-ray outburst observed with *ROSAT* by a slight *reduction* in the accretion rate (close to the Eddington limit), decreasing the WD photospheric radius on a hydrodynamic time-scale and accounting for the temperature change from an unobservable ≤ 20 eV to the observed 40 eV.

Further *ROSAT* monitoring of RXJ0513 from December 1992 to December 1993 (during the AO-3 period) found the source in an X-ray off-state 4 times, and in an on-state in July 1993 at a brightness comparable to that in November 1990 (Hasinger, 1994; Schaeidt, 1996). The *ROSAT* PSPC lightcurve showing the latter 5 observations, as well as the earlier ones discussed by Schaeidt et al. (1993), is shown in Fig. 3.24. This made RXJ0513 the first SSS observed to exhibit repeated X-ray outbursts on time-scales of years.

Considering the X-ray outburst as a nuclear shell flash on the WD surface, and by using the relation between M_{wd} and the critical envelope mass presented by Fujimoto (1982), Kahabka (1995) used the recurrence time of the X-ray outburst to estimate the mass of the WD to be in the $(1.33\text{-}1.38) M_{\odot}$ range, just under the Chandrasekhar limit (see the discussion in §2.6.4).

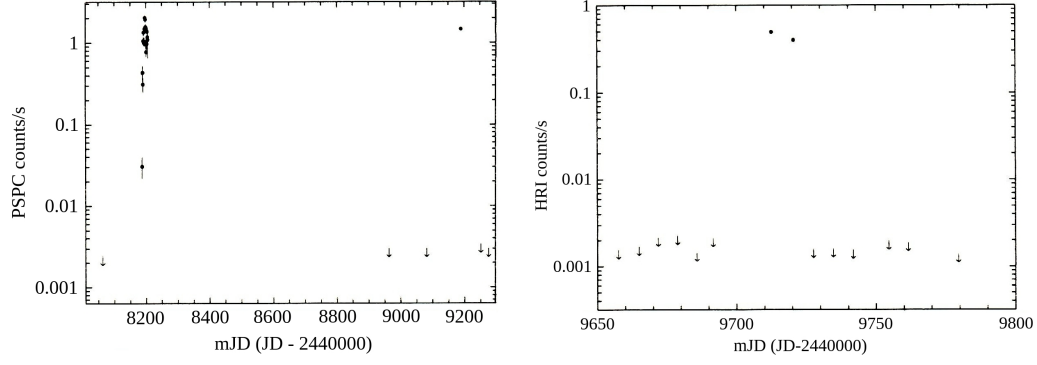


Fig. 3.24: *ROSAT* lightcurves of RXJ0513. Left: PSPC lightcurve from July 1990 to November 1993. Two on-states were observed, and the off-state observations are indicated with their upper limits. Right: HRI lightcurve from November 1994 to March 1995, showing an X-ray on-state lasting more than 2 weeks, with the X-ray turn-off occurring within ~ 7 d. (Adopted from Schaeidt, 1996, Fig. 1 and 2.)

Blackbody models were fitted to the first two X-ray on-states observed with *ROSAT* (Schaeidt et al., 1993; Schaeidt, 1996). The column densities were consistent with LMC membership. Reinsch, van Teeseling, Beuermann and Abbott (1996) fitted LTE WD model atmospheres to the July 1993 (AO-3) spectrum. Gänsicke et al. (1998) fitted an LTE model to the same spectrum. From a re-analysis of the *ROSAT* PSPC data with line-blanketed LTE WD atmosphere models, and a comparison to theoretical relationships between the properties of WDs with hydrogen shell burning, Suleimanov and Ibragimov (2003) obtained new estimates for the source properties of RXJ0513, and estimated the WD mass to be in the $0.8\text{--}1.2 M_{\odot}$ range. The best fit parameters for all these models are provided in Table 3.6, together with other X-ray model parameters for RXJ0513 from the literature.

RXJ0513 was also monitored with the *ROSAT* HRI from November 1994 to March 1995, and a third X-ray on-state was observed at the end of December 1994 (Fig. 3.24). From this lightcurve, Schaeidt (1996) constrained the time-scale of the X-ray turn-off to ~ 7 d, similar to the turn-on time observed in the survey data. The X-ray on-state lasts for $\sim 2\text{--}4$ weeks. These time-scales are too short to be explained by the Van den Heuvel et al. (1992) limit cycle model or the recurrent SSS model of Kahabka (1995).

Reinsch et al. (1999) and Reinsch et al. (2000) presented additional *ROSAT* data, where the source was once again observed in outburst. They monitored RXJ0513 with the HRI through one complete outburst that started at the end of August 1997, as shown in Fig. 3.30 (which will only be presented in §3.2.3, where the correlation with the MACHO lightcurve will be discussed). The X-ray flux increased by a factor of >100 and reached a maximum within ~ 5 d. The outburst lasted ~ 40 d, and the X-ray flux declined again by a factor of >100 within ~ 2 d. During the outburst, the X-ray flux declined slowly with an exponential time constant of 34 ± 6 d. Reinsch

Table 3.6: A summary of the X-ray properties of RXJ0513, as obtained from published models fitted to the X-ray spectra. BB stands for blackbody and LTE refers to WD model atmosphere in local thermodynamic equilibrium, while non-LTE (NLTE) models allow departures from LTE. For more details regarding the particular WD atmosphere models, please refer to the relevant articles. Suleimanov and Ibragimov (2003) assumed the distance d to the source to be 50 kpc (this was not explicitly stated by the other authors, but we assume a similar value was used). Where available, the χ^2 -value and degrees of freedom (dof) of the fit are also given. Where only a single value is given in the last column, it represents the reduced χ^2 -value.

Facility	Model	$\log g$ (cgs)	kT_{eff} (eV)	N_{H} (10^{20} cm^{-2})	R (10^8 cm)	L_{bol} (10^{37} erg s^{-1})	χ^2 /dof
<i>ROSAT</i> (survey) ^[1,2]	BB	–	39 ± 7	9	~ 28	23.5 ± 7.8	
<i>ROSAT</i> (AO-3) ^[2]	BB	–	37 ± 2	11			
<i>ROSAT</i> (AO-3) ^[3]	LTE		50	7	10^{a}	9.5	
<i>ROSAT</i> ^[4]	LTE	8		4-7	6-12	2.5-9	
<i>ROSAT</i> ^[4]	LTE	9			4-9		
<i>ROSAT</i> ^[5]	LTE ^b	8.0-8.5	50.4 ± 0.9	7.24^{a}	8-12	5.5-9.5	33.8/9
<i>XMM-Newton</i> ^[6]	BB	–	$43.95^{+0.55}_{-3.45}$	$6.2^{+0.3}_{-0.1}$		10-20	>1.9
<i>Chandra</i> ^[7]	BB	–	$43.5^{+2.2}_{-6.6}$	$7.11^{+0.03}_{-0.63}$	25.3	29.4	643/180
<i>Chandra</i> ^[7]	LTE ^b	8.5	40.1	7^{a}	13.2	5.75	1170/181
<i>Chandra</i> ^[7]	LTE ^{b,c}	8.5	33.2	7^{a}	20	7.5	859/179
			52.7				

[1] Schaeidt et al. (1993)

[2] Schaeidt (1996)

[3] Reinsch, van Teeseling, Beuermann and Abbott (1996)

[4] Gänsicke et al. (1998)

[5] Suleimanov and Ibragimov (2003)

[6] McGowan et al. (2005, 2006)

[7] Burwitz et al. (2007)

^a Fixed during fit.

^b Line-blanketed.

^c Two-temperature model: the cooler component encompassing 96.5% of the emitting surface, and the hotter component 3.5%.

et al. (2000) pointed out significant variations of ± 0.1 HRI counts s^{-1} on time-scales of hours to days, but did not find any periodicity in these variations.

RXJ0513 was observed by *XMM-Newton* 8 times in the April-May 2004 period (McGowan et al., 2005, 2006). Simultaneous modelling of all the EPIC pn spectra with an absorbed blackbody yielded the parameters given in Table 3.6. The spectral parameters also varied significantly between the observations, as will be further discussed in §3.2.3. Several blueshifted absorption features were visible in the *XMM-Newton* RGS spectra, indicating a high-velocity (up to $\sim 3000 \text{ km s}^{-1}$) outflow consisting of ionized gas (see Fig. 3.25). This could be associated with the hot regions of the bipolar jet that is known to exist in the source, or possibly the expansion of the WD envelope itself, as highly ionized by-products of nucleosynthesis stream out from the WD surface.

Burwitz et al. (2007) reported on the first high-resolution *Chandra* LETGS spectrum of RXJ0513 that was recorded on 24 December 2003, just after the source entered an optical low state. These authors fitted 3 different models to the data: (i) a simple blackbody model, (ii) a single temperature line-blanketed LTE WD model atmosphere, and (iii) a two-temperature line-blanketed LTE WD model atmosphere. The fit parameters are also listed in Table 3.6. The spectrum is shown in Fig. 3.26, together with a single temperature WD model spectrum.

The reason for fitting a two-temperature spectrum was that a secondary minimum in the χ^2 value was found at a higher temperature. The lower temperature fits the longer wavelength part of the spectrum ($> 40 \text{ \AA}$) well, while the short-wavelength portion is fitted well by the higher temperature model. The short wavelengths may be modified strongly by a wind from the WD, and by radiation from the disc corona, therefore the longer wavelengths should possibly yield a more accurate reflection of the actual conditions close to the WD. It was noted that the derived surface area of the cool component is ~ 25 times larger than the area of the hot component. Another possibility is that the WD envelope could be structured, with a hot component partially covered by cooler material.

3.2.2 Ultraviolet

Pakull et al. (1993) obtained UV spectra of RXJ0513 from the *IUE* on 24 and 25 April 1992. The spectrum recorded with the SWP camera is shown in Fig. 3.27. The spectrum is dominated by the emission lines of He II $\lambda 1640$ (with a red wing extending to $\sim 1670 \text{ \AA}$), N V $\lambda\lambda 1239, 1243$ and interstellar absorption lines. The Si IV $\lambda\lambda 1397, 1402$ emission is hardly visible. Surprisingly, the C IV $\lambda\lambda 1548, 1551$

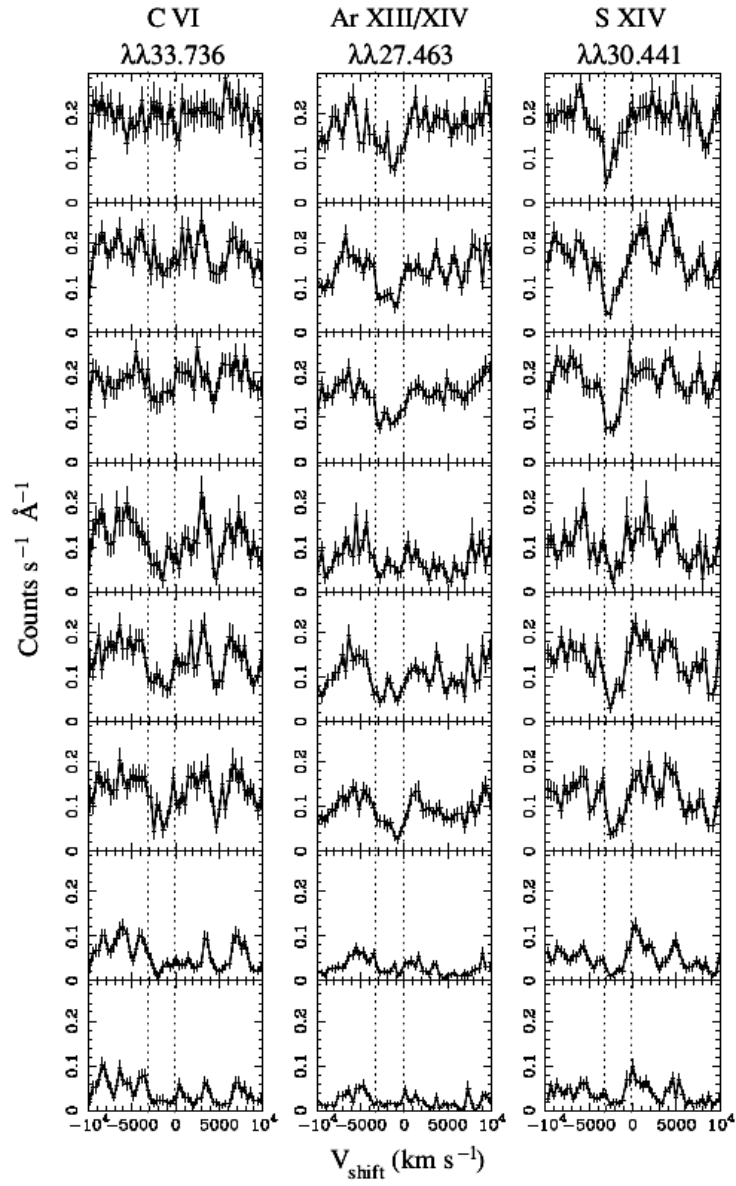


Fig. 3.25: Absorption features in the *XMM-Newton* RGS spectra of RXJ0513 during April-May 2004. The features indicate outflow velocities of up to ~ 3000 km s⁻¹. (Adopted from McGowan et al., 2005, Fig. 7.)

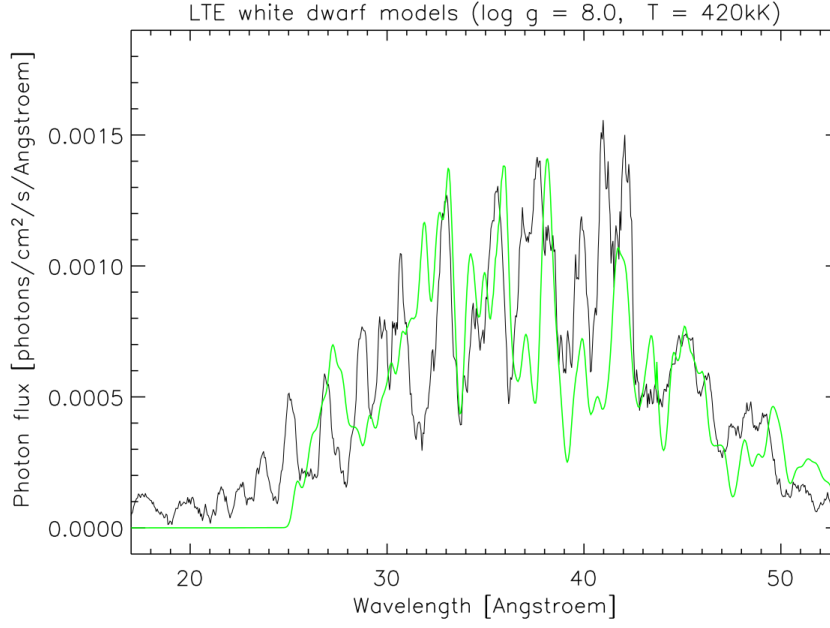


Fig. 3.26: The *Chandra* LETGS spectrum of RXJ0513 on 24 December 2003 (black), with an absorbed LTE WD model atmosphere with $\log g = 8$, $kT_{\text{eff}} = 36$ eV and $N_{\text{H}} = 6.5 \times 10^{20} \text{ cm}^{-2}$ overplotted (green). (Adopted from Burwitz et al., 2007, Fig. 3.)

Table 3.7: The equivalent widths of emission lines in the *IUE* spectrum of RXJ0513 in April 1992 (Pakull et al., 1993).

Line ID	Equivalent width (Å)
He II $\lambda 1640$	−10
N V $\lambda \lambda 1238, 1240$	−4
Si IV $\lambda \lambda 1397, 1402$	Weaker than −0.8
C IV $\lambda \lambda 1548, 1551$	Weaker than −1.5

emission, which is usually the strongest feature in the UV spectrum of CVs, LMXBs and most other hot stars, is extremely weak in the RXJ0513 spectrum. The equivalent widths of the emission features are given in Table 3.7.

RXJ0513 was observed with the GHRs on *HST* on 13 November 1996 (Gänsicke et al., 1998,b). The dominant emission features superimposed on the almost flat continuum are the N V $\lambda \lambda 1239, 1243$ resonance doublet and the O V $\lambda 1371$ emission line (Fig. 3.28). Faint O IV $\lambda 1345$ may be present. There are several interstellar absorption lines in the spectrum as well. An additional unidentified absorption feature is present at 1227.2 Å. If it is considered as redshifted Ly α absorption in an outflow within the source, it corresponds to a velocity of approximately +2800 km s $^{-1}$. However, the formation of an absorbed *red-shifted* component in an outflow would be difficult to explain.

The N V doublet lines are centred on a systemic velocity of approximately +300 km s $^{-1}$, and have a FWHM of ~ 1 Å. Both of them seem to have a double-peaked structure,

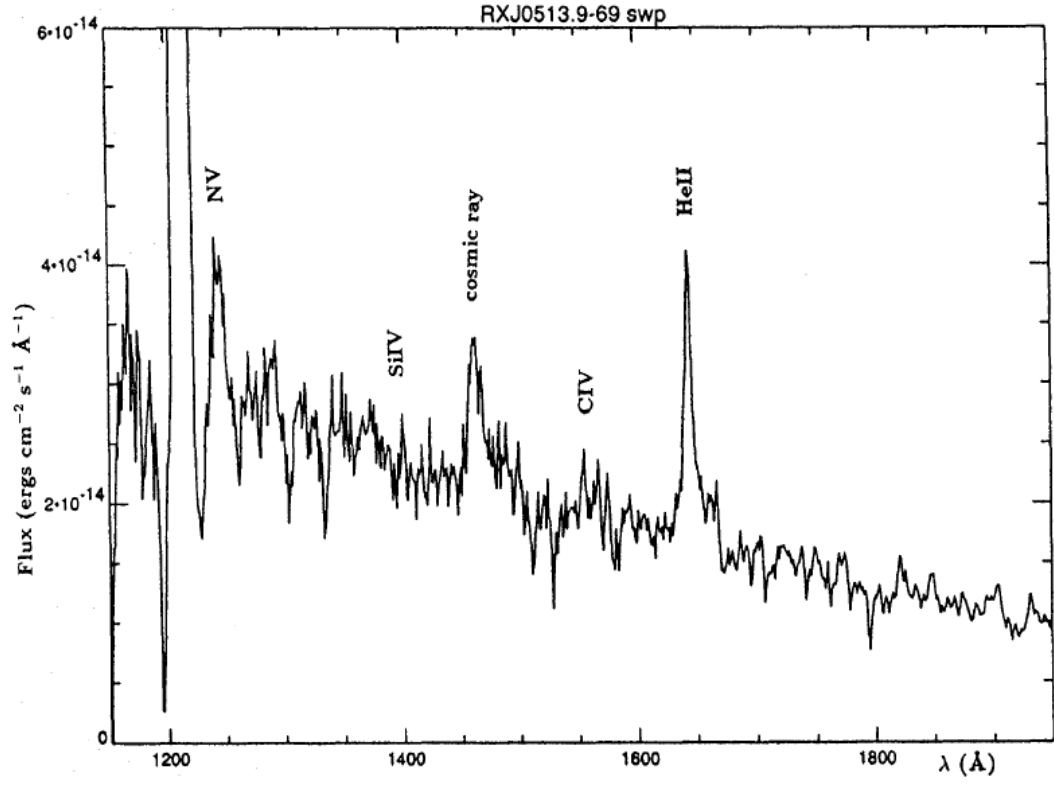


Fig. 3.27: The *IUE* SWP spectrum of RXJ0513 on 24 April 1990. (Adopted from Pakull et al., 1993, Fig. 2.)

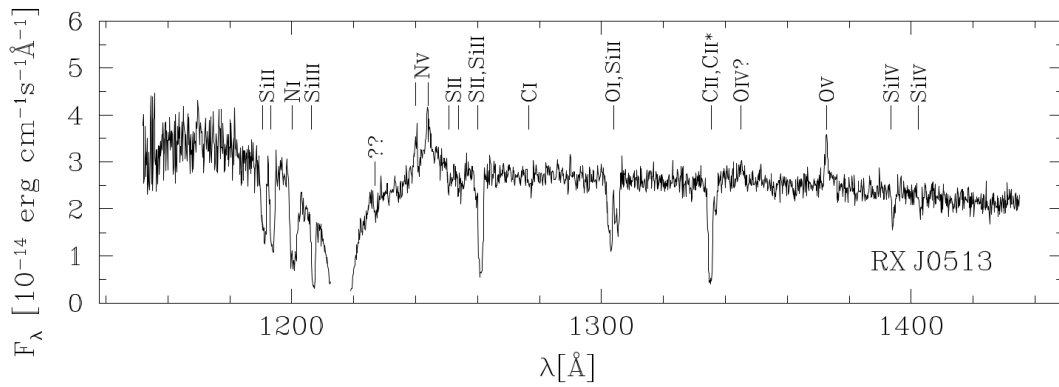


Fig. 3.28: The *HST* spectrum of RXJ0513 on 13 November 1996. The geocoronal emission line has been truncated. (Adopted from Gänsicke et al., 1998, Fig. 1.)

and an additional red wing extending to approximately $+800 \text{ km s}^{-1}$ from the central wavelength. The narrow component of O v $\lambda 1371$ also has $\text{FWHM} \sim 0.9 \text{ \AA}$, and a red extension to $+800 \text{ km s}^{-1}$. The small widths of the narrow components of these lines, as well as the apparent double peaks in N v, are consistent with an irradiated accretion disc observed at a low inclination (Gänsicke et al., 1998).

By modelling the broad Ly α absorption feature, Gänsicke et al. (1998) determined an upper limit of $N_{\text{H}} = (5.5 \pm 1.0) \times 10^{20} \text{ cm}^{-2}$ for the hydrogen column density between our solar system and RXJ0513. This includes interstellar absorption in the Galaxy and the LMC, and also any intrinsic absorption taking place in the binary system itself, and is consistent with the value obtained during their LTE model atmosphere fits to *ROSAT* data (see §3.2.1).

In a similar manner as was done for CAL 83 in §3.1.2, one can compare this value with the galactic foreground column density of $6.5 \times 10^{20} \text{ cm}^{-2}$ according to Dickey and Lockman (1990) and Zimmermann et al. (1994), or the $\sim (3-8) \times 10^{20} \text{ cm}^{-2}$ according to Fitzpatrick (1986) and Bessell (1991). The total neutral hydrogen column density *within* the LMC along the line of sight towards RXJ0513 is $\sim 6 \times 10^{20} \text{ cm}^{-2}$ (Rohlfs et al., 1984). One can therefore conclude that RXJ0513 is located at the near side of the LMC, and that there is no substantial intrinsic neutral hydrogen absorption in the source.

Gänsicke et al. (1998) also analysed the *IUE* data of RXJ0513: two observations reported on by Pakull et al. (1993), and an additional spectrum obtained on 29 June 1995. The flux in the June 1995 spectrum was a factor of ~ 2 lower than the April 1992 spectra, while the flux in the *HST* spectrum was comparable to the latter. We note that the June 1995 observation was performed shortly after the optical decline at $\text{JD} \sim 2448900$ commenced (see Fig. 3.29), which could explain the lower UV flux. Gänsicke et al. (1998) remarked that the emission at $1666 \text{ \AA}$ may represent He II $\lambda 1640$ emission redshifted by approximately $+4000 \text{ km s}^{-1}$.

Hutchings et al. (2002) obtained *FUSE* observations of RXJ0513 during a 24 h period on 30 October 2001. The spectrum with a range of $916-1182 \text{ \AA}$ was characterized by a hot continuum, narrow O VI $\lambda\lambda 1032, 1038$ emission superimposed on very broad O VI emission with P Cyg absorption, narrow He II $\lambda 1085$ emission and weak Lyman absorption. N III and C III emission was not detected, even though it is present in the optical spectra. Although the narrow lines exhibited small radial velocity variations similar to that observed for the optical lines (see §3.2.3), the radial velocity curves of the broad O VI emission and Lyman absorption have much larger semi-amplitudes of 117 ± 40 and $54 \pm 10 \text{ km s}^{-1}$ respectively.

Consequently, Hutchings et al. (2002) proposed a scenario where $\sim 120 \text{ km s}^{-1}$ represents an upper limit to the motion of the WD, and thus a high value for the mass function. They considered the lack of an eclipse in the photometric lightcurve, the high projected jet velocities observed in the optical, together with their findings from the *FUSE* data, and consequently proposed an orbital inclination of $\sim 35^\circ$, a WD mass of $\sim 0.8 M_\odot$ and an evolved secondary with a mass of $1.6 M_\odot$ and luminosity class III. Although these proposed parameters are not definite, it is noted that the higher mass function is in good agreement with the model proposed by Van den Heuvel et al. (1992).

Additional *FUSE* observations on 5 July 2003 exhibited characteristics similar to that of the 2001 spectra (Hutchings et al., 2005, 2006). A P Cyg absorption structure was also noted on the blue side of the He II $\lambda 1085$ line. However, the radial velocity modulation of the broad O VI emission has a semi-amplitude of $\sim 26 \text{ km s}^{-1}$, much smaller than the $\sim 117 \text{ km s}^{-1}$ measured in 2001. This is an indication that the 2001 modulation, or possibly both, was not dominated by orbital motion but rather represented mass flows within the system. The broad O VI emission ($\sim 2500 \text{ km s}^{-1}$) may represent the rotational velocity of gas orbiting near the WD surface.

3.2.3 Optical

Photometry

Upon inspection of the ESO Schmidt UBVR plates, Pakull et al. (1993) identified an object with a “UV excess” as the optical counterpart of RXJ0513. The ICRS coordinates of the source are 05:13:50.812 –69:51:47.67.

From direct *V*- and *B*-band imaging with the 2.2-m Max-Planck/ESO telescope in April 1992, Pakull et al. (1993) reported small but significant night-to-night variability between $V = 16.6$ and 16.8 . No variability was found in the colour index $B - V = -0.14 \pm 0.04$. Pakull et al. also considered the optical magnitude of the system from ESO Schmidt plates taken for the MACHO project during October to December 1990. Although the MACHO observations did not exactly coincide with the *ROSAT* observations, Pakull et al. inferred that the optical flux did not increase during the X-ray outburst; on the contrary, it seemed to fade slightly from $B = 16.5$ to 17 .

Cowley et al. (1993) confirmed the optical counterpart identification of Pakull et al. (1993). It corresponds to the star identified as Harvard Variable 5682 by Leavitt (1908). By considering measurements of RXJ0513 since the early 20th century, Cowley et al. found that the optical magnitude is indeed variable, but that it maintained a *V*-magnitude of roughly 16.5 ± 0.5 over the past century. They also

obtained B - and V -band photometry of the source at the CTIO in December 1992 and March 1993. They determined mean magnitudes of $V = 16.769 \pm 0.044$ and $B - V = -0.112 \pm 0.012$ (with a possible systematic error of up to 0.05 mag in the latter), with the V -band measurements varying from 16.684 to 16.814. Cowley et al. (1996) obtained additional photometric measurements in December 1993, and pointed out that the source exhibits ± 0.1 mag night-to-night variations, and ± 0.3 mag variations on a time-scale of months.

Crampton et al. (1996) also reported on new photometric data obtained with the CTIO 0.9-m telescope and the SAAO 1.0-m telescope on 3-6 November 1994. During the first and last nights, the source was “bright”, with $V = 16.837 \pm 0.012$, but declined to a “faint” state with $V \sim 17.0$ -17.1 on the night of 4-5 November. These authors also considered all the photometric data from 1992 to 1994, but could not find any orbital modulation in the photometric lightcurve, indicating a small inclination angle. The V -magnitude of the system was 16.78 in December 1992, 16.99 in 1993, and 16.84 in 1994.

Reinsch, van Teeseling, Beuermann and Abbott (1996) conducted an optical multi-colour campaign with the Dutch 0.9-m telescope at ESO between October 1994 and March 1995, obtaining measurements on an approximately weekly basis in the Bessel B -, V - and R -filters (see also Reinsch, van Teeseling, Beuermann and Thomas, 1996). During the optical high state, the V -magnitude variation was less than 0.3 mag. However, it declined by ~ 0.9 mag within ~ 10 days in December 1994, then slowly returned to the high state after ~ 30 days, remaining at a value ~ 0.2 mag higher than the pre-minimum magnitude for ~ 50 d. The decline was associated with a reddening of the optical spectrum, but Cowley et al. (2002) later remarked that this may be due to the contribution of a nearby field star. The optical low state coincided with an X-ray outburst observed with *ROSAT* in December 1994 (see §3.2.1), but it was not clear whether the X-ray increase occurred simultaneously with or after the optical decline.

In addition, Reinsch, van Teeseling, Beuermann and Abbott (1996) considered V - and B -filter magnitudes obtained as the by-product of spectroscopic observations at the ESO/MPI 2.2-m telescope between December 1992 and February 1994, and observed a similar decline event in December 1993. Another decline occurred in February 1993, albeit with sparse observational coverage.

Reinsch, van Teeseling, Beuermann and Abbott (1996) favoured the interpretation mentioned earlier, offered by Pakull et al. (1993) and supported by the calculations of Kato (1985), namely that the X-ray outbursts are caused by a contraction and associated increase in effective temperature of the WD envelope. In the optical high

state, the accretion rate approaches the Eddington rate, and the WD photosphere expands considerably. The lower effective temperature at almost constant luminosity causes the X-ray flux to drop below the observable limit, and the optical flux to be enhanced. Additionally, the larger WD photospheric radius enhances the amount of disc illumination, increasing the reprocessed optical flux and line emission from the disc even more. The expansion also causes an increased rate of mass loss from the WD (Kato, 1985), explaining the prominent P Cyg profile during the optical high state (see the next section for a discussion of spectroscopic properties). A slight decrease in the accretion rate causes the photosphere to contract and the WD once again becomes an SSS, entering the optical low state.

Reinsch, van Teeseling, Beuermann and Abbott (1996) considered the results of their X-ray high-state LTE model atmosphere fit (see §3.2.1), estimated the X-ray luminosity during the X-ray low/optical high state, and then calculated, for the optical low *and* high states, the total expected *V*-magnitude of the accreting WD, the irradiated accretion disc and the unheated and heated sides of the secondary, as well as the *B* – *V* colour of the disc. Their results fit their observations remarkably well, weighing in favour of the contracting/expanding WD photosphere model.

Fortunately, RXJ0513 was also included in a field that was frequently monitored by the 1.27-m telescope at Mount Stromlo Observatory in Australia from 22 August 1992 to 27 November 1995, as part of the MACHO project. This lightcurve is shown in Fig. 3.29. The system was found to exhibit significant optical variability on time-scales of ~ 100 -200 d, cycling between optical low and optical high states (Alcock et al., 1996). The transition from optical high to optical low involves a brightness decrease of ~ 0.8 -1.0 mag within ~ 10 d. Then, the brightness usually increases with approximately 0.3-0.4 mag within ~ 8 d after the transition, and maintains this level before making a rapid transition back into the high state after remaining in the low state for ~ 40 d. The magnitude during the high states is not constant, but can exhibit erratic modulations of up to ~ 0.3 mag, and also an overall fading of ~ 0.2 mag from the start to the end of a high state.

Alcock et al. (1996) derived a photometric orbital period of 0.76278 ± 0.00005 d from the optical high state measurements of the first 3 years of MACHO data, which is in good agreement with the spectroscopic period derived by Crampton et al. (1996) (see the next section). The folded lightcurve is sinusoidal in shape, with a semi-amplitude of 0.0213 ± 0.0009 mag, once again indicating a low inclination. By considering the Van den Heuvel et al. (1992) model, Alcock et al. estimated the density of the secondary to be consistent with a ~ 2.5 -3.0 $M_{\odot}$ main sequence star of spectral type $\sim A0$.

The orbital period of Alcock et al. (1996) is quite close to the 0.745 ± 0.015 d derived

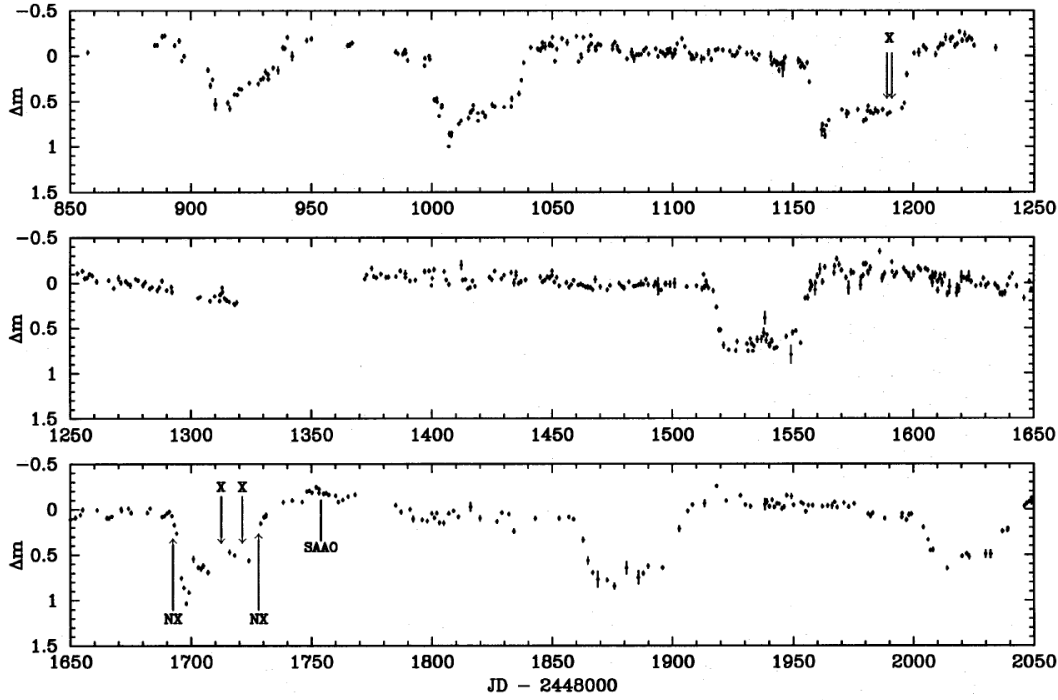


Fig. 3.29: The MACHO lightcurve of RXJ0513 from August 1992 to November 1995, taken in the “blue” filter, which is approximately equivalent to the Johnson V -band. The times of X-ray on-states (X) and off-states (NX) are indicated, as well as the epoch of the SAAO 1.0-m fast optical photometry. (Adopted from Southwell et al., 1996a, Fig. 3.)

by Motch and Pakull (1996) from photometric measurements with the 0.9-m Dutch telescope from 15 to 20 December 1992, during an optical high state with $B = 16.58$. However, the latter lightcurve seemed to exhibit a rather narrow minimum and not a sinusoidal form, and a ~ 0.1 mag semi-amplitude.

Southwell et al. (1996a) plotted the MACHO lightcurve first published by Alcock et al. (1996), and indicated the epochs of observed X-ray on- and off-states (Fig. 3.29). It is clear that the X-ray outbursts all occurred during optical low states, while the X-ray off-states were observed during optical high states. This was a very important result, which strongly suggests an anti-correlation between the X-ray and optical flux. Southwell et al. (1996a) favours the expansion and contraction of the WD photosphere discussed by Kato (1985), Pakull et al. (1993) and Reinsch, van Teeseling, Beuermann and Abbott (1996) as the cause for this behaviour, and presented a comprehensive model of photospheric contraction.

They proposed that the accretion rate is very high ($\sim 10^{-6} M_{\odot} \text{ yr}^{-1}$) during the “normal”, optical high state, with the WD atmosphere slightly inflated, probably not more than $3R_{\text{wd}}$. The accretion rate now decreases by a factor of 3 or more, possibly because of a magnetic spot on the secondary covering the L_1 point, as suggested for VY Scl stars by Livio and Pringle (1994). The lower accretion rate causes

a contraction of the WD photosphere, shifting the emission peak from the ultraviolet to soft X-rays and causing the X-ray outburst. It is noted that this would require the X-ray outburst to be *preceded* by the optical decline, which is reconcilable with the data in Fig. 3.29.

The X-ray flux from the WD may now irradiate the donor and cause an increase in the mass transfer rate, either because the X-ray irradiation inflates the material above the donor’s photosphere, or because it heats up the cooler magnetic spot area. This might also explain the “overshoot” of ~ 0.2 mag when the source enters the optical high state again, because the attenuation of the mass transfer rate may cause it to be temporarily higher directly after recovery than directly before the attenuation. As mentioned by Koyama et al. (1999), the well-developed accretion disc during the optical high state may also play a role in shielding the secondary from some of the irradiation by the WD, which could in turn contribute to the lower accretion rate that introduces an optical low state. Southwell et al. (1996a) showed that the photosphere of the WD would be able to contract on the observed time-scale of ~ 7 d if the WD is massive ($\sim 1.3 M_{\odot}$).

This model is supported by the findings of Meyer-Hofmeister et al. (1997), who investigated the structure of the accretion disc rim in SSSs (see §2.4.5). They showed that the optical magnitude of RXJ0513 in the optical low and high states can be reproduced by modelling the changing irradiation of the accretion disc rim as the radius of the WD envelope changes.

Southwell et al. (1996a) also remarked that the main difference between RXJ0513 and CAL 83 is that the ratio of the accretion rate to the rate of nuclear burning is probably higher in the former, a notion that is supported by the higher optical luminosity of RXJ0513. It is interesting to note that, if the accretion rate is indeed $\sim 10^{-6} M_{\odot} \text{ yr}^{-1}$ in RXJ0513, the accretion disc would be stable against irradiation-induced warping (Southwell et al., 1997, see also the discussion in §2.4.4 of this thesis), which is not the case in CAL 83, where the disc is probably precessing as a result of warping.

From photometric observations at the SAAO 1.0-m telescope, on 5 February 1995, Southwell et al. (1996a) found significant optical variability (~ 0.1 mag) on time-scales $\lesssim 3$ h, although no periodicities were uncovered (upper limit ~ 0.015 mag). Photometric measurements by Cowley et al. (1998) indicated a source magnitude of $V = 16.810 \pm 0.007$ in November 1995.

At the distance of the LMC, and assuming an average reddening of $E(B - V) = 0.1$ mag, the absolute visual magnitude of RXJ0513 during the optical high state is

$M_V \sim -2$ (Pakull et al., 1993; Cowley et al., 1993; Crampton et al., 1996; Alcock et al., 1996; Southwell et al., 1996a). This is much brighter than most accretion-driven LMXBs (van Paradijs and McClintock, 1995), and typical values of $M_V = +4$ to $+7$ found in CVs (van Paradijs, 1983). Alcock et al. (1996) applied the Warner (1987) empirical relation between the orbital period and the absolute magnitude of the accretion disc for dwarf novae discs at maximum, and obtained $M_V = +1.1$ for the measured $P_{\text{orb}} \sim 0.76$ d, much fainter still than the measured M_V . This argues in favour of nuclear burning as the dominant energy source in addition to the accretion energy released. For $M_{\text{wd}} \sim M_{\odot}$ and $i \sim 10^\circ$, Eq. (2.44) indicates an accretion rate of $\sim 10^{-5} M_{\odot} \text{ yr}^{-1}$, of the order of the Eddington value. If a large fraction of the optical luminosity arises from irradiation of the disc by the nuclear burning WD, such an extreme accretion rate would not be required.

Reinsch et al. (2000) shifted the MACHO lightcurves in time so that the start of the optical decline occurs at $\Delta t = 0$. Comparison of MACHO data taken almost simultaneously with their HRI monitoring of a complete X-ray outburst (see §3.2.1) indicated that the X-ray outburst commenced within a day or two after the onset of the decline in the optical. The correlation of the HRI and MACHO lightcurves are shown in Fig. 3.30.

Reinsch et al. (2000) used LTE WD model atmosphere spectra to calculate the dependence of the photospheric radius on the HRI count rate. They used the binary lightcurve code BINARY++ (van Teeseling, 1998) to determine the optical magnitude as a function of photospheric radius, considering the contribution of the WD, as well as the irradiated disc and companion star. Their predicted lightcurve is superimposed on the MACHO measurements in Fig. 3.30, and fit the optical lightcurve remarkably well.

The photospheric expansion/contraction model for the cycles exhibited by RXJ0513 is supported by Reinsch et al. (2000), but they presented a detailed, self-sustained limit cycle model based on this concept that does not require an external accretion rate modulation like the occurrence of star spots. Their model is shown in Fig. 3.31. Four different time-scales characterize the long-term behaviour of the source: the ~ 140 d the source spends in the X-ray low, optical high state, followed by a rapid transition ($\lesssim 4$ d) into the X-ray high, optical low state, where it remains for ~ 30 d, and finally transitions to the X-ray low, optical high state again within $\lesssim 2$ d.

At the (arbitrary) start of the cycle in Fig. 3.31(d), an accretion disc supplies material to a non-expanded WD at a rate close to the Eddington rate. The WD photosphere starts expanding (Fig. 3.31(d)-(e)), increasing the amount of disc irradiation. The increased disc temperature enhances the disc viscosity and therefore increases

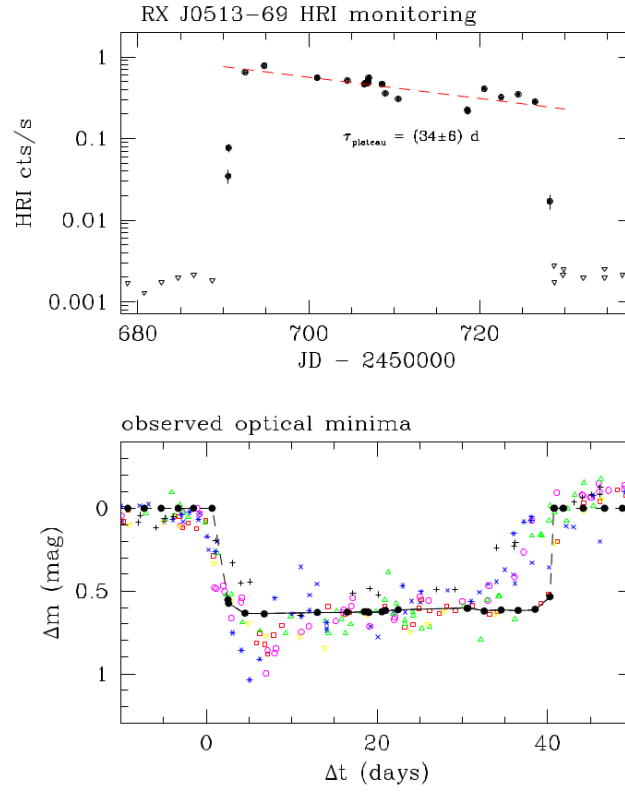


Fig. 3.30: The correlation between the X-ray and optical flux of RXJ0513. The MACHO lightcurves during different optical low states have been overplotted in the bottom panel. The filled circles in the bottom panel represents the predicted optical lightcurve from the modelling of Reinsch et al. (2000). (Adopted from Reinsch et al., 2000, Fig. 1.)

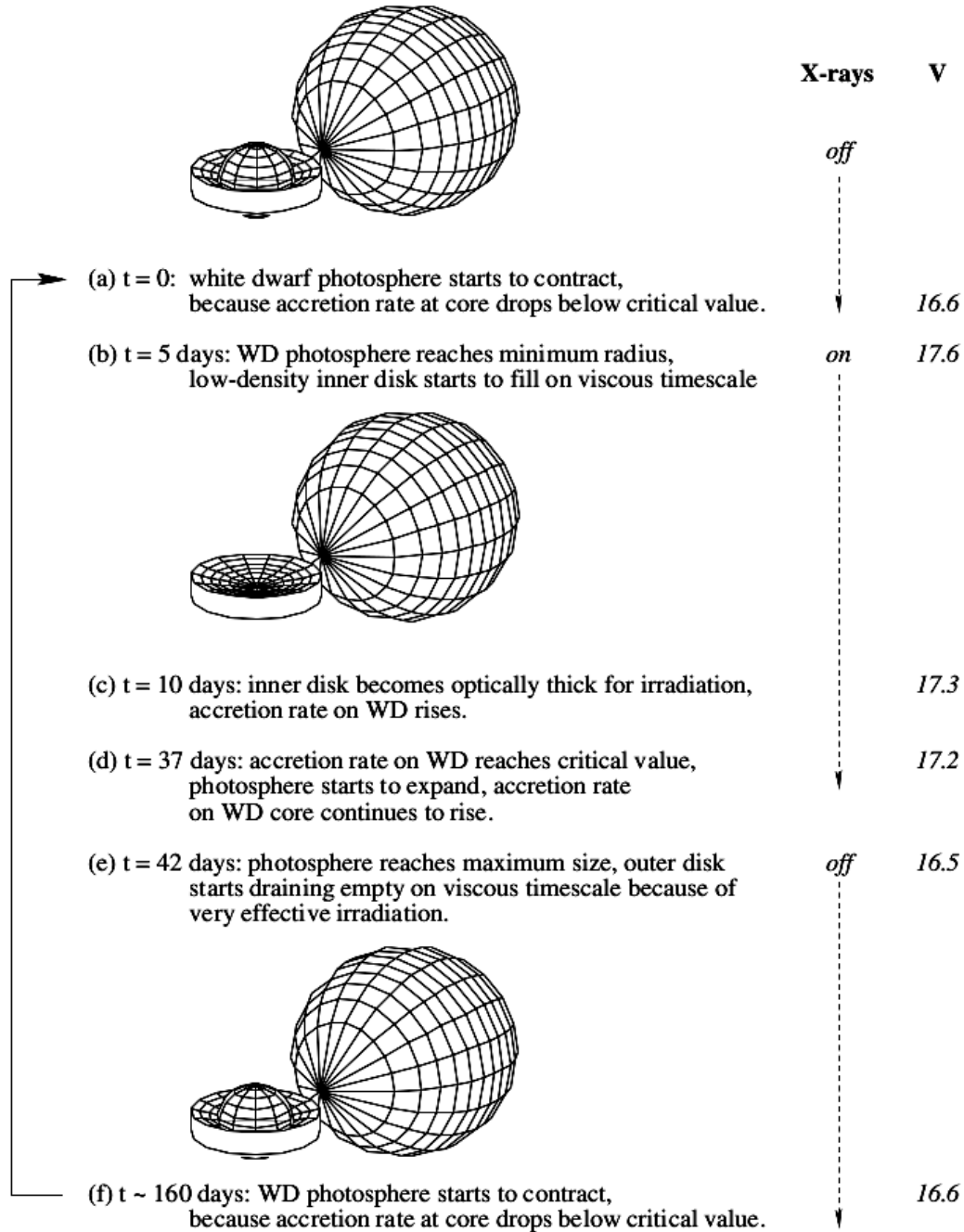


Fig. 3.31: The limit cycle model proposed by Reinsch et al. (2000) for RXJ0513. (Adopted from Reinsch et al., 2000, Fig. 3.)

the rate of mass flow through the disc so that it exceeds the mass transfer rate from the donor, and the mass in the disc is drained onto the WD at a viscous time-scale estimated at ~ 130 d (Fig. 3.31(e)). A consequent decrease in the accretion rate causes the WD photosphere to contract (Fig. 3.31(f), (a)), and the collapsing envelope leaves a “hole” in the inner accretion disc with approximately the envelope radius. The accretion disc gradually fills in the hole on a viscous time-scale of ~ 40 d (Fig. 3.31(a)-(d)), and the cycle starts over.

It was noted that the expansion/contraction of the WD photosphere needs to be by a factor of $\gtrsim 7$, to yield a decrease/increase in the effective temperature by a factor of $\gtrsim 3$. The corresponding changes in X-ray luminosity and optical magnitude is shown in the right hand side of Fig. 3.31.

Cowley et al. (2002) presented the extended MACHO lightcurve recorded from 1992 to 1999 (see Fig. 3.33 later in this section). They noted significant variability in the spacing, depths and duration of the optical minima. The separation between minima ranged between 104 and 195 d, averaging at ~ 161 d. The average depth is ~ 0.64 mag, ranging between approximately 0.3 and 1.2 mag. The duration of the minima increased somewhat from ~ 35 to ~ 50 d from the start to the end of the MACHO observations. No significant optical colour changes could be found between the optical high and low states. From the optical high state observations, Cowley et al. derived an orbital ephemeris of

$$T_{\min} = \text{JD } 2448858.099 \pm 0.001 + (0.7629434 \pm 0.0000020)E \text{ days}, \quad (3.2)$$

By folding both the “blue” and “red” MACHO bands through this orbital period, it was shown that (i) the two colours exhibit the same modulation through the binary orbit, i.e. the orbital modulation is not associated with a colour change, and (ii) orbital modulation with the same phasing and amplitude can be found in the optical low state. Cowley et al. also presented new photometric measurements obtained with the CTIO 0.9-m telescope on 21-28 March 1999, that are completely consistent with the MACHO results.

After removing the orbital modulation from the high state optical data, Cowley et al. (2002) discovered a significant optical periodicity with $P = 83.2$ d and a full amplitude of ~ 0.03 mag that was not present during the optical low states. They speculated that this period may arise from disc precession. If this is indeed the case, we note that this may imply a somewhat lower mass transfer rate than usually assumed for RXJ0513, in order to make it unstable to warping (see §2.4.4).

Hachisu and Kato (2003*a,b*) presented an alternative limit cycle model to the one

proposed by Reinsch et al. (2000). They show that, once the mass accretion rate onto the WD exceeds the critical rate of $\sim 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$, the excess material is blown off in an optically thick wind from the expanding WD (as discussed in §2.6.3). The wind mass loss rate can increase from zero to $\sim 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$ within a day, obscuring the soft X-rays and introducing the X-ray low, optical high state. The disc is now in an extended state. The winds collide with the donor, stripping off its outer layer, with the stripped gas being lost from the system, and thus attenuating the mass transfer. The mass in the disc is drained by accretion onto the WD and the disc shrinks to its original size, while the WD photosphere also shrinks gradually. Once the accretion rate drops below the critical rate, the accretion wind ceases to blow within about a day, revealing the underlying X-ray emission, and the source enters the X-ray high, optical low state again.

Hachisu and Kato (2003b) showed that this limit cycle can explain the extremely rapid turn-on and turn-off time-scales more readily than the limit cycle of Reinsch et al. (2000). Consideration of their modelled lightcurves implies that the WD in RXJ0513 is massive, with a mass of 1.2-1.3 $\text{M}_{\odot}$.

McGowan et al. (2005) reported new photometric measurements obtained with the SAAO 1.9-m telescope, the CTIO 1.3-m telescope and the Barfold 14-in during the December 2002 to June 2004 time period (see also McGowan et al., 2006; Charles et al., 2010). Their *XMM-Newton* observations (see §3.2.1) were triggered during the optical low state that started in April 2004. By fixing the column density to the value obtained for the combined fit, they fitted a blackbody model to each individual EPIC pn spectrum. They found that the X-ray temperature and luminosity decreased during the optical low state. The photospheric radius initially decreased, but started to increase again as the optical flux started to climb again, with a total radius expansion by a factor of 3.9. These observations strongly support the contracting/expanding WD photosphere model discussed previously: the effective temperature reaches a maximum near the beginning of the X-ray on-state, and decreases as the envelope extends again.

Simultaneous UV measurements with the *XMM-Newton* Optical Monitor indicates that the optical and UV flux are roughly correlated, but the UV flux starts to increase somewhat before the optical at the end of the optical low state. This strongly favours the expansion of the WD photosphere during the transition from the optical low to the optical high state, as the emission peak gradually shifts from the soft X-ray region into the ultraviolet.

Burwitz et al. (2008) obtained new photometric data with the SMARTS/CTIO 1.3-m telescope, and the CTIO 1.0-m telescope from August 2003 to January 2004, and

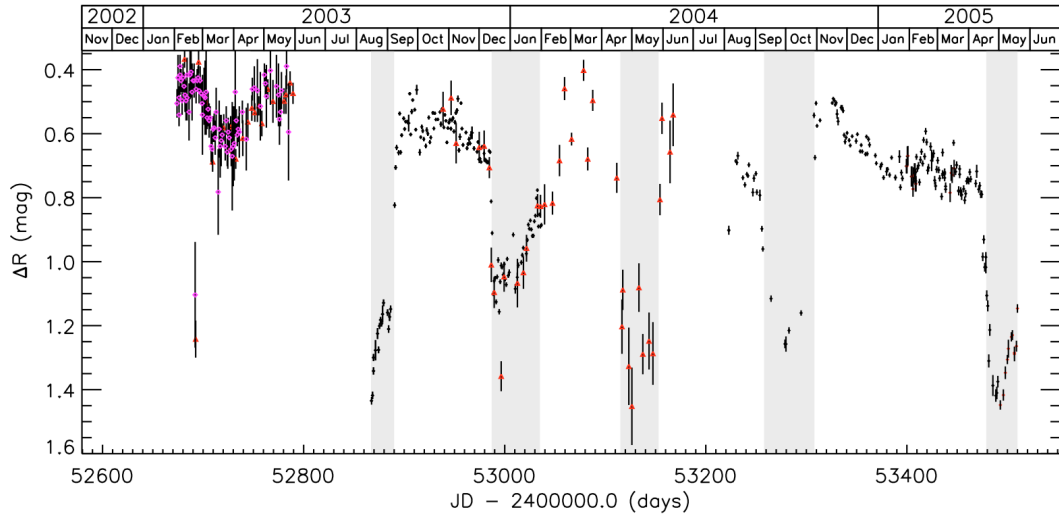


Fig. 3.32: Optical monitoring reported by Burwitz et al. (2008) (black crosses), together with that of McGowan et al. (2005) (red triangles for Barfold Observatory and pink diamonds for CTIO 1.3-m). The optical low states are shaded in grey. (Adopted from Burwitz et al., 2008, Fig. 1.)

from August 2004 to May 2005. These data, together with the photometry presented by McGowan et al. (2005) is shown in Fig. 3.32. Burwitz et al. (2008) scrutinized these measurements, together with the MACHO lightcurve (Cowley et al., 2002), and found that the duration of the optical high state varies between 63 and 171 days. They employed the relation between cycle length and mass transfer derived by Hachisu and Kato (2003*a,b*), and found that the mass transfer rate varied in the $\sim(2.5-13) \times 10^{-6} M_{\odot} \text{ yr}^{-1}$ range on time-scales of years (i.e. by a factor of ~ 5). Burwitz et al. (2008) also illustrated the existence of a stable stationary solution where the accretion rate is able to adjust to the wind from the photosphere, preventing the natural occurrence of the accretion wind limit cycle. They concluded that a yet unknown hysteresis is required to explain the cyclic behaviour of the source.

RXJ0513 was also monitored with the 1.27-m telescope at Mount Stromlo Observatory as part of the OGLE project (phase III). Rajoelimanana et al. (2013) investigated the long-term variability of the source by analysing the MACHO and OGLE-III lightcurves, which are shown in Fig. 3.33. By performing a Lomb-Scargle period analysis of these lightcurves, they derived a quasi-period of 168 ± 1 d for the long-term variability (Fig. 3.34). The durations of the optical low and high states are ~ 30 and ~ 138 d, corresponding to absolute visual magnitudes M_V of approximately -1.3 and -2.4 respectively. Rajoelimanana et al. (2013) also derived an orbital period of 0.762956 ± 0.000005 d from the OGLE-III lightcurve (Fig. 3.35).

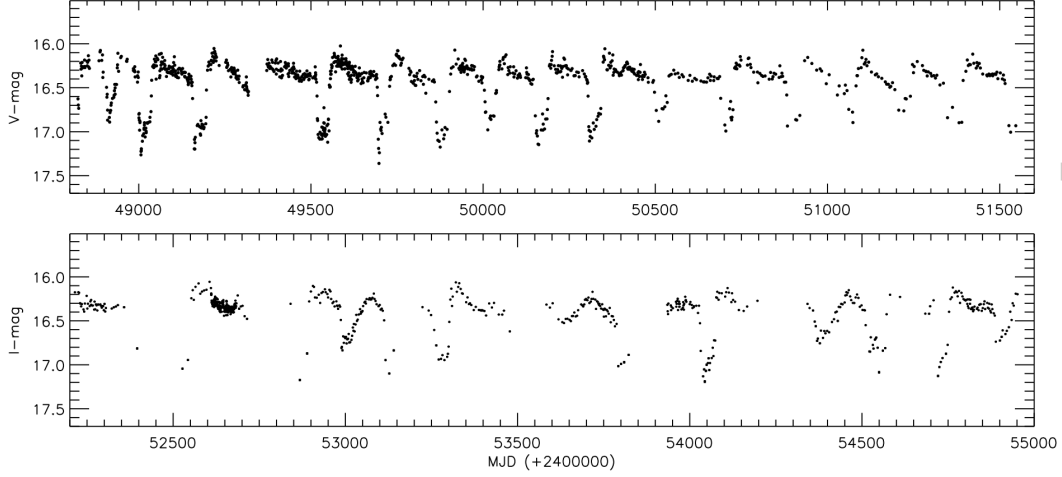


Fig. 3.33: The MACHO (top) and OGLE-III (bottom) lightcurves of RXJ0513. The quasi-periodic cycles recur every ~ 168 d, with the optical high states lasting ~ 140 d. (Adopted from Rajoelimanana et al., 2013, Fig. 2)

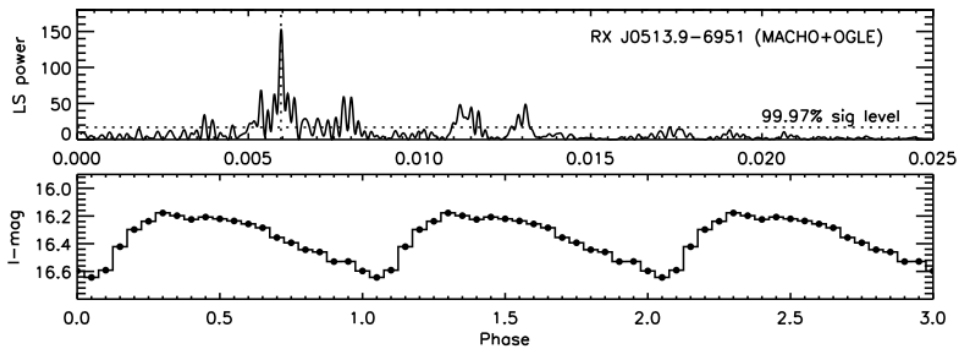


Fig. 3.34: The Lomb-Scargle (LS) periodogram (top) and folded MACHO + OGLE-III lightcurve (bottom) of RXJ0513, illustrating the long-term ~ 168 d period. (Adopted from Rajoelimanana et al., 2013, Fig. 3.)

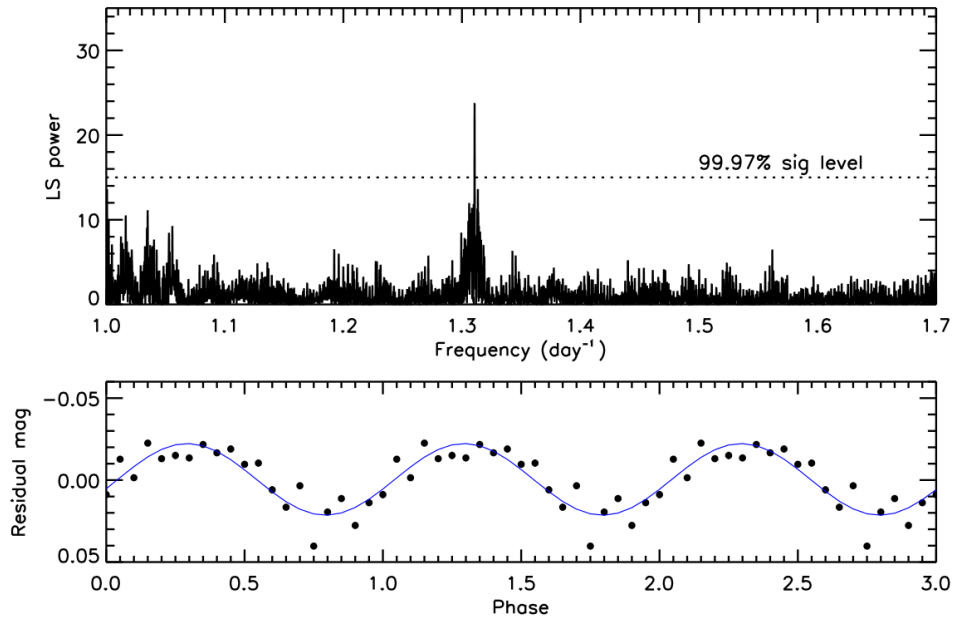


Fig. 3.35: The Lomb-Scargle (LS) periodogram (top) and folded OGLE-III lightcurve (bottom) of RXJ0513, illustrating the derived orbital period of 0.762956 ± 0.000005 d. The phasing is arbitrary. (Adopted from Rajoelimanana et al., 2013, Fig. 5)

Spectroscopy

Optical spectra obtained by Pakull et al. (1993) with the 2.2-m Max-Planck/ESO telescope from 20 to 25 April 1992, exhibit Balmer, He II and O VI emission, and are very similar to the spectrum of CAL 83. Blending of He II lines with the Balmer emission is expected. The He II $\lambda 4686$, H β , H γ and H δ emission lines consist of a narrow component superimposed on a significantly broader component, which has a FWHM of ~ 30 Å for He II $\lambda 4686$ (see Fig. 3.36). The He II $\lambda 4686$ and higher-order Balmer lines also exhibit P Cyg profiles, with the edge of the blue absorption component at a velocity of ~ 3800 km s $^{-1}$. These indicate outflows with velocities comparable to the escape velocity of a WD.

Equivalent widths measured by Pakull et al. (1993) are given in Table 3.8. These authors also remarked on the high temperatures in the emission region as indicated by the presence of O VI and the absence of He I emission. The narrow line components exhibited night-to-night radial velocity variations between 190 and 300 km s $^{-1}$, presumably due to binary orbital motion, with the mean value close to the systemic velocity of the LMC.

Pakull et al. (1993) found that the UV-to-optical continuum exhibits a $F_\nu \propto \nu^{1.5 \pm 0.4}$ shape, and has a very high luminosity that should be associated with a highly luminous accretion disc and a high mass transfer rate. These authors also considered the LMXB model suggested by Greiner et al. (1991), with the luminosity derived

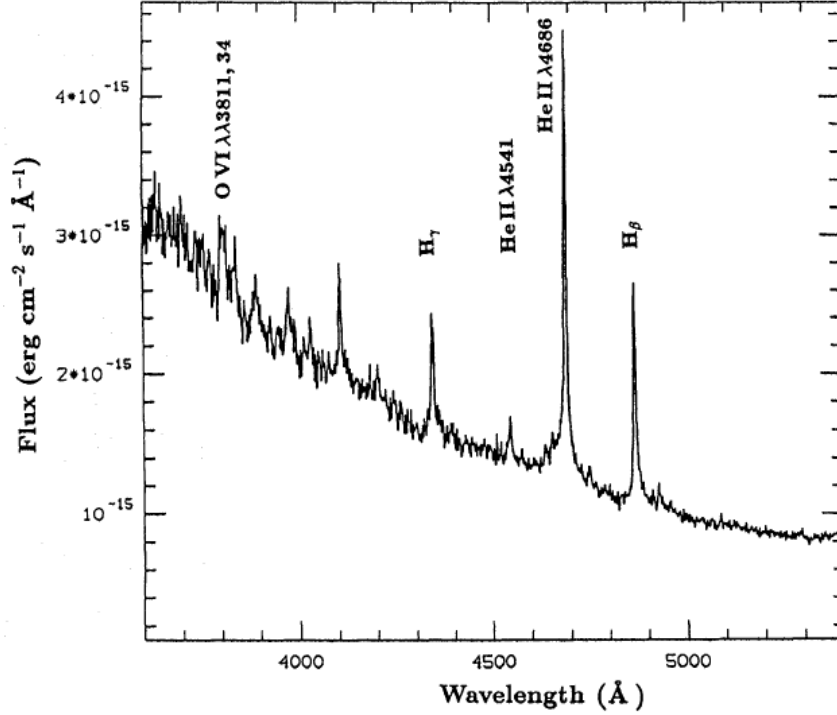


Fig. 3.36: The mean spectrum of RXJ0513 obtained with the EFOSC 2 imaging spectrograph. The line profiles of He II $\lambda 4686$, H β , H γ and H δ consist of a narrow component superimposed on a significantly broader component. The Balmer lines exhibit weak blueshifted absorption. (Adopted from Pakull et al., 1993, Fig. 3.)

Table 3.8: The equivalent widths of emission lines in the 2.2-m Max-Planck/ESO telescope spectrum of RXJ0513 in April 1992 (Pakull et al., 1993).

Line ID	Equivalent width (Å)
O VI $\lambda\lambda 3811,34$	-4
H β +He II $\lambda 4859$	-16
He II $\lambda 4686$	-33 (narrow+broad component)

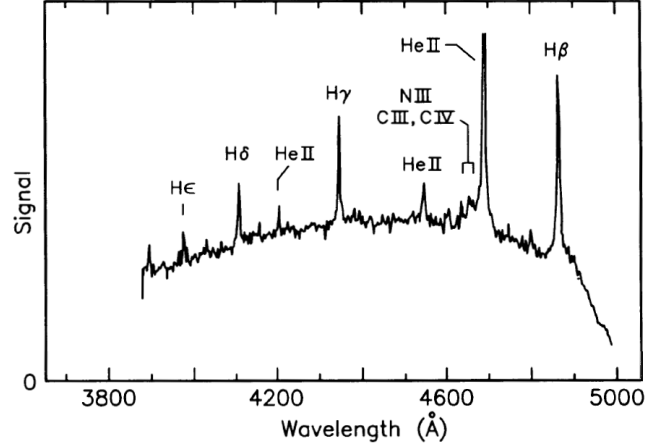


Fig. 3.37: The average spectrum of RXJ0513 obtained in December 1992 with the CTIO 4-m telescope. The He II $\lambda 4686$ line has been truncated. (Adopted from Cowley et al., 1993, Fig. 2.)

from accretion onto a neutron star. However, they remarked on the fact that such a high accretion rate onto a neutron star with a magnetic moment of 10^{30} G cm³ would imply an Alfvén radius 10 times smaller, hence a blackbody temperature >3 times higher than that of the observed spectrum.

Cowley et al. (1993) also observed RXJ0513 with the Cassegrain spectrograph on the CTIO 4-m telescope in December 1992. The mean spectrum is shown in Fig. 3.37, and exhibits strong He II and Balmer emission, as well as the N III, C III and C IV blend at ~ 4640 Å. The emission lines have strong narrow components superimposed on broad base components. The base of the He II $\lambda 4686$ line has a FWHM of ~ 25 Å, and the base of H β a full width at zero intensity of ~ 20 Å. The Balmer lines are noted to be asymmetrical, with weak blue absorption. The radial velocity range of the narrow peaks is $\sim 50 \pm 15$ km s⁻¹. The average He II velocity is $+292$ km s⁻¹, consistent with LMC membership, with the velocity of the Balmer lines being ~ 35 km s⁻¹ larger due to the blue absorption. The equivalent width of the He II $\lambda 4686$ emission is approximately -16 Å, and for H β it is -10 Å. Small variations may exist in the equivalent widths.

Cowley et al. (1996) reported on 15 spectra obtained at the CTIO during December 1993. Considering the nightly radial velocity variations of ~ 40 km s⁻¹, an orbital period of 0.43 d was suggested, although alias periods at 0.32 and 0.77 days were also noted. These authors also reported on several weak, unidentified lines in most of their spectra, which could be explained as highly Doppler shifted components of H I and He II emission at velocities of approximately $+4200$ and -3700 km s⁻¹. These so-called “satellite” features were proposed to be due to a bipolar outflow or “jets”. This notion was confirmed by their private communication with Pakull (1994), who also reached this conclusion while analysing independent data.

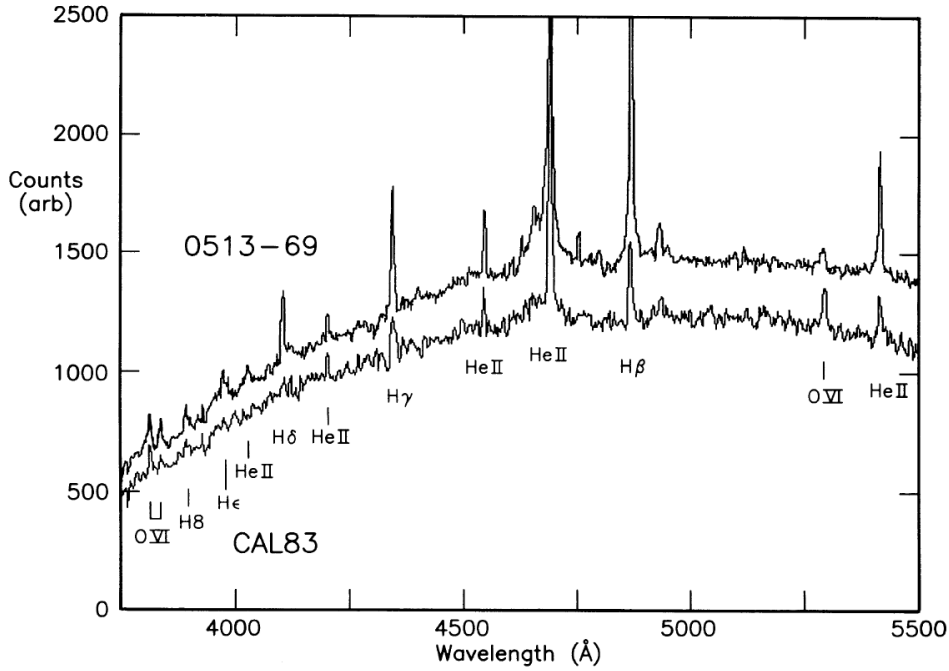


Fig. 3.38: The average spectrum of RXJ0513 obtained with the CTIO 4-m telescope in 1994, shown with the spectrum of CAL 83 for comparison. (Adopted from Crampton et al., 1996, Fig. 2.)

Crampton et al. (1996) obtained spectroscopic datasets of RXJ0513 with the CTIO 4-m telescope during 1992-1994. Their average low-resolution spectrum of 1994 is shown in Fig. 3.38, together with a spectrum of CAL 83 they recorded for comparison. The spectra are dominated by He II emission and Balmer + He II blends, with bipolar blueshifted and redshifted high-velocity components at $\sim 4000 \text{ km s}^{-1}$. High ionization O VI emission is present at 3811, 3834 and 5290 Å, with probable weak O V emission near 5595 Å. Emission of N V $\lambda\lambda 4603, 4619$ is also visible. Although C IV emission was deficient in the UV spectrum (Pakull et al., 1993), weak C IV $\lambda\lambda 5801, 12$ emission and possibly $\lambda 4658$ (blended with the blue wing of He II $\lambda 4686$) appear in the optical spectrum. Even though C III emission at 5695 Å is absent, another C III multiplet at 4650 Å seems to contribute to the He II $\lambda 4686$ wing as well. A line near 4070 Å is probably also C III emission. The average equivalent widths of the main emission lines are given in Table 3.9.

The equivalent widths of the He II emission lines were $\sim 50\%$ weaker in 1993 than in 1992 and 1994, and the Balmer lines $\sim 40\%$ weaker. The blue wing of He II $\lambda 4686$ did not appear to vary, i.e. the wing itself and/or the contributing ionized lines did not change as much as the narrow Balmer and He II components. No significant modulations in line shapes and intensities were observed on shorter time-scales of a few days.

Table 3.9: The average equivalent widths of the emission lines of RXJ0513 measured with the CTIO 4-m telescope during 1992-1994. Note that the values should be interpreted as being negative, as these are emission features. (Adopted from Crampton et al., 1996, Table 3).

Wavelength (Å)	ID	1992	1993	1994
4340	H γ	4.8	1.2	3.3
4861	H β	13.1	3.5	8.8
4200	He II	0.6	0.3	0.6
4541	He II	1.1	0.9	1.3
4686	He II	22.	9.4	19.
5411	He II	...	2.1	2.9
4624	S $^-$	1.1	0.6	0.7
4748	S $^+$	0.8	0.5	0.7
4795	S $^-$	0.4	0.3	0.3
4927	S $^+$	...	0.8	0.8

A close-up view of the He II λ 4686, H δ and H β emission is shown in Fig. 3.39. The blueshifted and redshifted satellite features are indicated with S $^-$ and S $^+$ respectively, and the possible ions contributing to the emission are also shown. The average spectra in the region containing He II λ 4686 and H β with their satellite features are shown in Fig. 3.40 for 1992 up to 1994. For 1994, both the low- and high-resolution spectra are given. After correcting the wavelengths of the satellite features for the systemic velocity, their wavelengths are 4623.6, 4748.1, 4795.5 and 4927.1 Å during 1993-1994, consistent with an average velocity of 4020 km s $^{-1}$ relative to the central component. However, the satellite features in the 1992 spectra (although these are noisier) seem to be $\sim$ 250 km s $^{-1}$ closer to the central line. Also, the He II λ 4686 satellites seemed to become extended towards lower absolute velocities on the last observing night in 1994. Possible jet features could also be present for H γ at 4404 and 4283 Å.

The FWHM for the He II λ 4686 and H β satellites is $\sim$ 6 Å, with the exception of S $^+$ of H β , which has FWHM $\sim$ 10 Å, and is probably blended with ionized emission. The equivalent widths are all approximately -0.7 Å, except for S $^-$ of H β , which was difficult to measure because of the blueshifted absorption. The variation in the equivalent widths of the central components did not seem to be accompanied by variations in the strengths of the satellite features.

By considering the radial velocity curves of the narrow He II and Balmer emission line components, Crampton et al. (1996) determined the most probable orbital period to be 0.77 d. The radial velocity semi-amplitude is $K_1 \sim 11 \pm 2$ km s $^{-1}$ around an average of +280 km s $^{-1}$, yielding a very small mass function and companion mass. It was illustrated by Crampton et al. (1996) that if the primary is a WD and *if* the measured semi-amplitude represents the WD orbital motion accurately, the donor needs to be evolved to be able to fill its Roche lobe. Also, the inclination

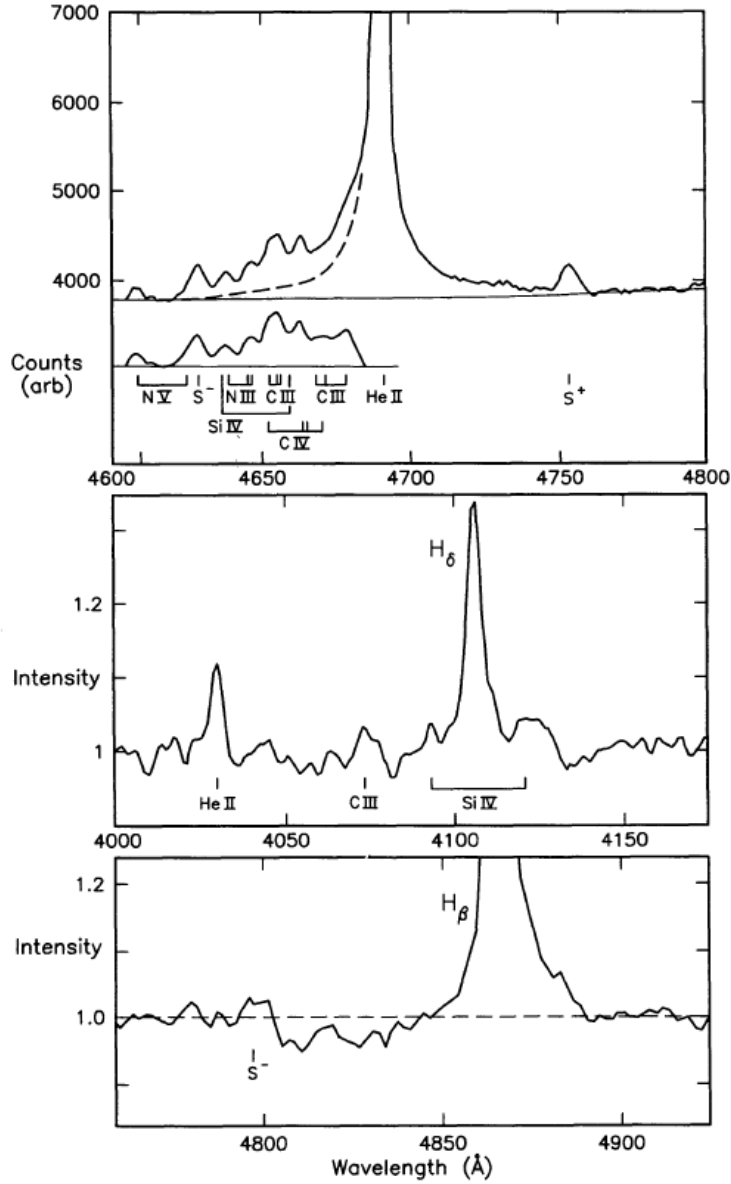


Fig. 3.39: The profiles of the He II $\lambda 4686$, H δ and H β emission lines observed in RXJ0513 with CTIO 4-m telescope. The jet satellite features are indicated, as well as possible contributing ions. The dashed line in the blue wing of He II $\lambda 4686$ represent a reflection of the weaker red wing to illustrate the amount of asymmetry. Note the P Cyg absorption profile of the H β line. (Adopted from Crampton et al., 1996, Fig. 7.)

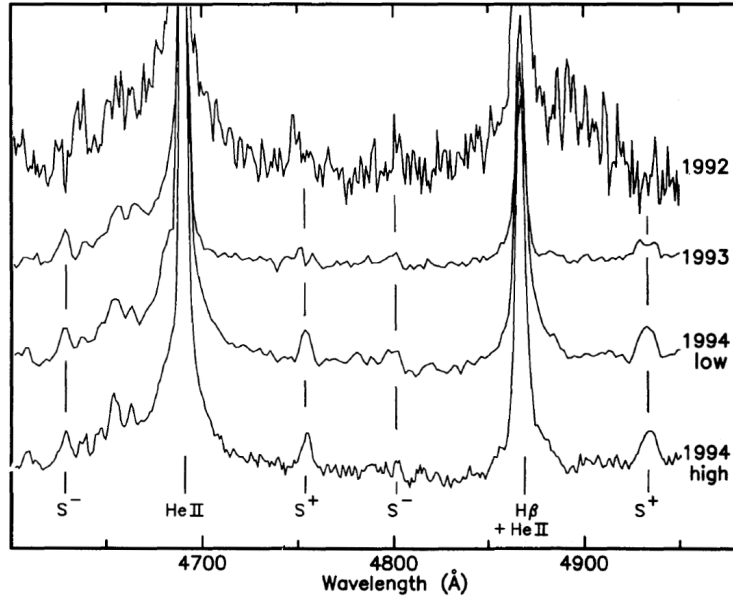


Fig. 3.40: An illustration of the variation of the He II $\lambda 4686$ and H β satellite features from 1992 to 1994. The displacement of these features seem to be slightly smaller in 1992. (Adopted from Crampton et al., 1996, Fig. 8.)

Table 3.10: The equivalent widths and fluxes of emission lines in RXJ0513 during the optical high (December 1992) and optical low (December 1993) states. The equivalent widths indicate negative values, as these are emission lines. (Adopted from Reinsch, van Teeseling, Beuermann and Abbott, 1996, Table 1.)

Line	optical high state		optical low state	
	EW (Å)	Flux (10^{-16} erg cm $^{-2}$ s $^{-1}$)	EW (Å)	Flux (10^{-16} erg cm $^{-2}$ s $^{-1}$)
H γ	5.1 ± 0.2	57 ± 2	1.0 ± 0.2	6 ± 1
He II 4541	1.8 ± 0.2	18 ± 2	1.0 ± 0.2	6 ± 1
He II 4686	29.0 ± 1.0	270 ± 10	7.0 ± 0.5	45 ± 3
H β	11.7 ± 0.2	102 ± 2	1.6 ± 0.2	10 ± 1

angle i should then be very small, meaning that we are viewing the WD close to a pole-on orientation.

Reinsch, van Teeseling, Beuermann and Abbott (1996) performed spectroscopic observations at the ESO/MPI 2.2-m telescope between December 1992 and February 1994. They remarked on the fact that the equivalent widths of the Balmer and He II emission lines decreased significantly during the optical low states (see Table 3.10). The P Cyg absorption features that extend to ~ 3000 km s $^{-1}$ during the optical high state, also became weaker.

Southwell et al. (1996a) obtained spectra of RXJ0513 with the 3.9-m Anglo-Australian Telescope (AAT) from 29 November to 1 December 1994 (see also Southwell et al.,

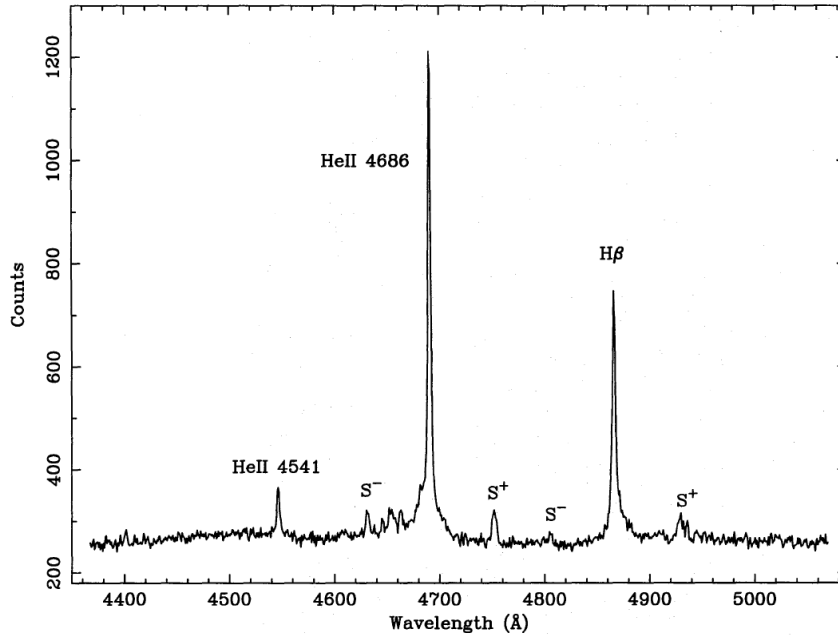


Fig. 3.41: The average blue spectrum of RXJ0513 observed with the AAT at the end of November 1994. (Adopted from Southwell et al., 1996a, Fig. 1.)

1996b). Their average blue and red spectra are shown in Fig. 3.41 and Fig. 3.42. They confirm the $H\beta$ and He II $\lambda 4686$ emission lines with narrow and broad components and satellite features observed previously, with N III, C III and C IV contributing to the extended He II $\lambda 4686$ wing, and marginal evidence of N V $\lambda\lambda 4603, 19$ emission. They also reported $\sim 3800 \text{ km s}^{-1}$ jet features in $H\alpha$. The equivalent widths of the major features are listed in Table 3.11. The line shapes and intensities did not change significantly over the 3 nights.

Southwell et al. (1996a) performed a combined radial velocity analysis of their data and that of Crampton et al. (1996), and derived a spectroscopic orbital period of $0.761 \pm 0.004 \text{ d}$ from the He II $\lambda 4686$ line, which is consistent with the result of Alcock

Table 3.11: The equivalent widths of emission lines in RXJ0513 as observed with the AAT. Values are given in Å, and should be interpreted as negative values, as these are emission features. (Adopted from Southwell et al., 1996a, Table 2.)

Line	Night 1	Night 2	Night 3	Overall Average
He II $\lambda 4541$	1.2 ± 0.3	1.5 ± 0.2	1.1 ± 0.1	1.4 ± 0.1
He II $\lambda 4686$	20.7 ± 0.8	19.2 ± 0.5	20.6 ± 0.2	20.4 ± 0.2
S_{4686}^-	0.9 ± 0.3	0.5 ± 0.2	0.9 ± 0.1	0.8 ± 0.1
S_{4686}^+	0.3 ± 0.3	1.0 ± 0.2	1.1 ± 0.1	1.1 ± 0.1
$H\beta$	13.2 ± 0.6	11.3 ± 0.4	10.7 ± 0.2	10.8 ± 0.1
$S_{H\beta}^-$	0.1 ± 0.3	0.0 ± 0.2	0.2 ± 0.1	0.2 ± 0.1
$S_{H\beta}^+$	1.3 ± 0.4	0.8 ± 0.3	1.4 ± 0.1	1.3 ± 0.1
$H\alpha$	48.2 ± 1.4	...	...	48.2 ± 1.4
$S_{H\alpha}^-$	1.6 ± 0.6	...	...	1.6 ± 0.6
$S_{H\alpha}^+$	2.6 ± 0.7	...	...	2.6 ± 0.7

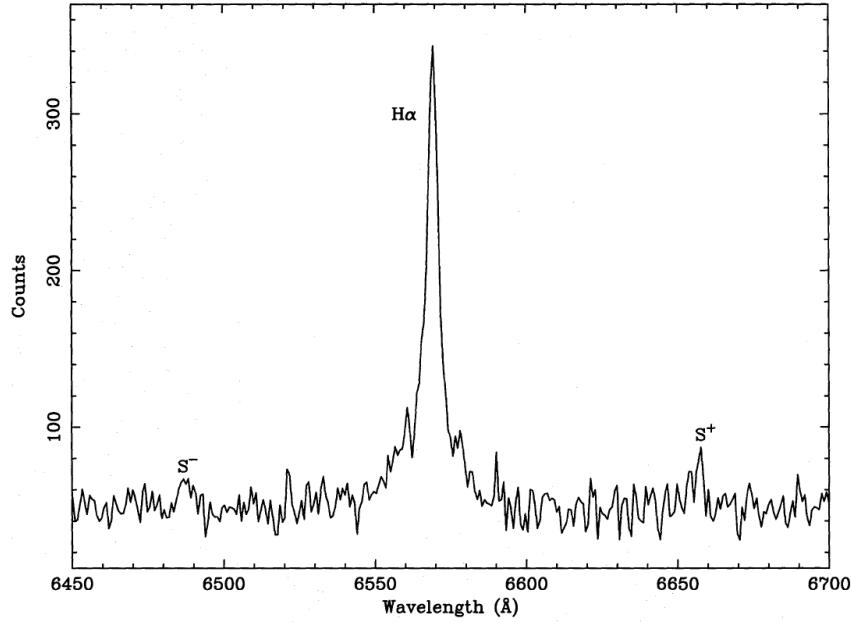


Fig. 3.42: The average red spectrum of RXJ0513 observed with the AAT at the end of November 1994. (Adopted from Southwell et al., 1996a, Fig. 2.)

et al. (1996). The semi-amplitude is $K_1 = 14.5 \pm 3 \text{ km s}^{-1}$ around a systemic velocity of $297 \pm 1 \text{ km s}^{-1}$. Southwell et al. confirmed that the very small mass function requires a very low inclination ($i \lesssim 7^\circ$) if the primary is a WD, and an evolved donor.

Spectroscopic observations of RXJ0513 were performed with the CTIO 4-m telescope from 3 to 6 November 1996 (Cowley et al., 1998). The RXJ0513 spectrum forms part of Fig. 3.14 that was included in §3.1.3. The spectra show similar characteristics to those described by previous authors. The average emission line measurements are given in Table 3.12. Cowley et al. (1998) compared the 1996 spectra to those obtained during an optical low state in 1993 and an optical high state in 1994. The main components of the Balmer and He II emission lines, as well as the red wings of He II $\lambda 4686$ and $H\beta$, and the C III, C IV, N III blend on the blue side of He II $\lambda 4686$, were all weaker in 1993 than in 1994 and 1996.

Cowley et al. (1998) also reported variability in the Balmer and He II satellite features (see Fig. 3.43 and Fig. 3.44). The S^- and S^+ features of He II $\lambda 4686$ and $H\beta$ were much weaker in 1996 than in 1993 and 1994, and may have lower velocities than was observed in 1994. The 1996 spectrum has an absorption component at 4809.8 Å , which was mostly in emission in 1993. The sharp emission feature at 4935.5 Å in the 1996 spectrum could be redshifted $H\beta$ emission, but the latter was much broader in 1993, indicating that it consists of a blend with other emission lines. Therefore the sharp feature in 1996 could be entirely due to an unidentified emission line contributing to the blend. The structures of the $H\alpha$ and $H\beta$ lines are similar.

Table 3.12: The average equivalent widths of the RXJ0513 emission lines as observed with the CTIO 4-m telescope in November 1996 (Cowley et al., 1998, Table 5). The Balmer lines may include contributions from the He II Pickering lines. The weaker O VI λ 3835 emission is not included in the O VI λ 3811 measurement.

Line ID	Measurement ($\text{\AA}$)
H α	-41.3
H β	-11.8
H γ	- 4.1
He II λ 4686	-22.5
He II λ 4541	- 1.5
He II λ 5411	- 3.1
O VI λ 5290	- 0.5
O VI λ 3811	- 1.7
FWHM of He II λ 4686	5.8

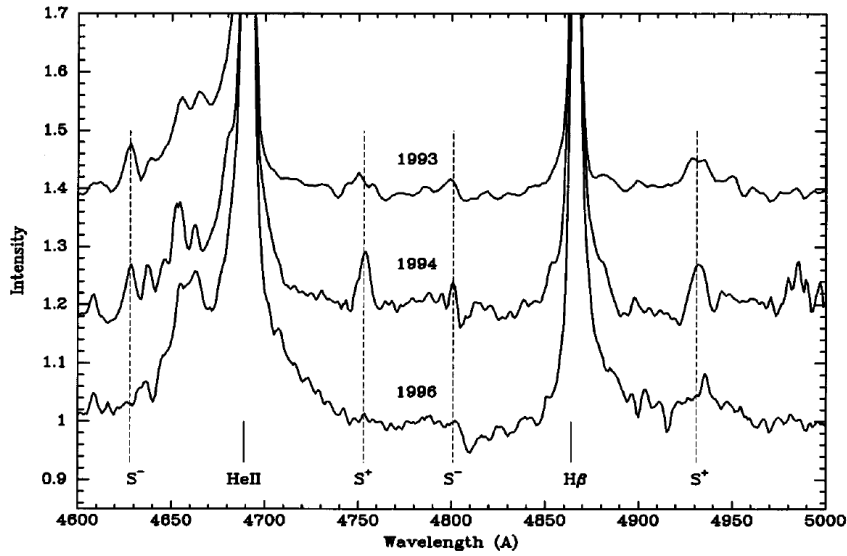


Fig. 3.43: The He II λ 4686 and H β emission from RXJ0513 in 1993, 1994 and 1996. (Adopted from Cowley et al., 1998, Fig. 3.)

The blueshifted Balmer absorption is at a velocity of -3332 km s^{-1} , comparable to that of the S^- feature of He II λ 4686. A feature at $\sim 6656 \text{ \AA}$ may be redshifted H α at a velocity of $+3976 \text{ km s}^{-1}$.

Cowley et al. (1998) remarked that the high-velocity P Cyg absorption indicates that the absorbed emission is *behind* the jet, which can be explained if our line of sight is almost parallel to the jet, i.e. the system is viewed at a low inclination. The variation in the shape and intensity of the absorption indicates a changing absorbing column or velocity profile, or perhaps a precession of the jet.

By using their updated orbital ephemeris (see the previous section), Cowley et al. (2002) re-investigated the He II λ 4686 radial velocity curves of Crampton et al. (1996) and Southwell et al. (1996a). It appeared that there might be a difference

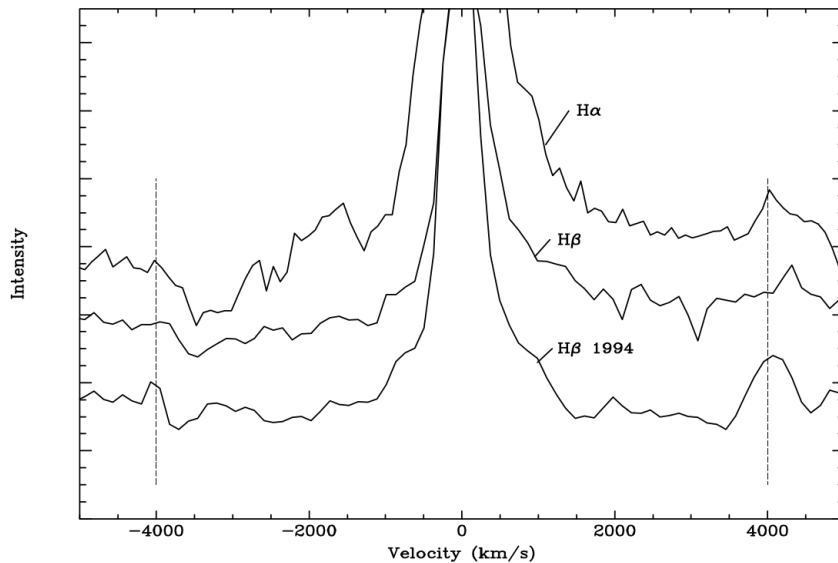


Fig. 3.44: A comparison of the $H\alpha$ and $H\beta$ emission and absorption in November 1996, and also $H\beta$ in 1994. The dashed lines at $\pm 4000 \text{ km s}^{-1}$ are for reference. (Adopted from Cowley et al., 1998, Fig. 4.)

in the phase, amplitude and average between the curve during optical high states and during the transition phase into a low state, indicating changes in the location of the dominant $\text{He II } \lambda 4686$ emission from the system. However, this result is not conclusive.

Remillard et al. (1995) performed a search for potential emission from an ionized nebula around RXJ0513, but did not find any extended $[\text{O III}]$ emission as was found in CAL 83. Crampton et al. (1996) confirmed the absence of extended $H\beta$, He II and $[\text{O III}]$ emission in their optical spectra.

3.2.4 Longer wavelengths

The ATCA observations of Fender et al. (1998) yielded an upper limit of 0.09 mJy for RXJ0513 at radio wavelengths of 3.5 cm and 6.3 cm.

3.3 MR Vel (RX J0925.7-4758)

3.3.1 X-ray

MR Vel was discovered as a Galactic supersoft X-ray source in *ROSAT* PSPC data obtained during the all-sky survey from July 1990 to January 1991 (Motch, Hasinger and Pietsch, 1994). The source was observed during 34 *ROSAT* orbits covering 3.5 d, with a total exposure time of 550 s. MR Vel exhibited X-ray variability of $\sim 50\%$ on time-scales of several hours. A pointed 3 ks *ROSAT* PSPC observation was performed on 20 November 1992, and the source exhibited no variability of $> 40\%$

on time-scales of seconds to minutes. A second pointed *ROSAT* observation was performed with the HRI on 23 November 1992. The X-ray intensity did not change by more than 50% between the two pointed observations.

The X-ray spectral energy distribution has a very narrow shape, only slightly larger than the instrumental response. A continuum model for the narrow spectrum requires a Wien exponential tail at higher energies and very steep interstellar absorption of the soft X-ray energies. From optical studies (see §3.3.2), Motch et al. (1994) constrained the column density in the direction of the source to $N_{\text{H}} = (1.0\text{--}1.9) \times 10^{22} \text{ cm}^{-2}$ (which is much larger than for the sources in the LMC). Their spectral analysis of the PSPC spectrum of 20 November 1992 showed that the spectrum can be fitted quite well by a blackbody model. The resulting blackbody parameters are given in Table 3.13, together with other X-ray fit results from the literature.

Hartmann and Heise (1997) fitted LTE and NLTE WD model atmospheres (without line opacities and line blanketing) to the same *ROSAT* data (see also Hartmann and Heise, 1996; Ebisawa, Asai, Dotani, Mukai and Smale, 1996). The fit parameters are listed in Table 3.13. These authors found a significant excess in the data above ~ 1.3 keV that is not fitted well by even the NLTE model. They pointed out that this is caused by the strong Ne IX absorption edge in the model that is not present in the data, and which cuts off the spectrum at 1.2 keV, and suggested that models with higher temperature and surface gravity, or simply more sophisticated models, should also be considered.

Hartmann and Heise (1997) also pointed out that all the spectral fits up to that time indicated a radius significantly smaller than that of a WD. However, it is noted that the upper limit of 2 kpc for the distance to MR Vel was based on an Eddington-limited blackbody model for a WD (Motch et al., 1994), while it is well-known that a blackbody model overestimates the bolometric WD luminosity. Thus, it may well be that the source is situated at a larger distance: at ~ 10 kpc one would obtain a photospheric radius comparable to that of a WD, and also larger bolometric luminosities (from the WD model atmosphere fits) more comparable to what is observed in other supersoft X-ray binaries. Alternatively, Hartmann and Heise suggest that the small derived radii may indicate soft X-ray emission from only a small fraction of the WD surface, i.e. a “hot spot” or “hot belt”, or that the WD is partly obscured by an accretion disc.

MR Vel was observed with *ASCA* on 22-23 December 1994 (Ebisawa et al., 2001, see also Ebisawa, Asai, Dotani, Mukai and Smale, 1996; Ebisawa, Asai, Mukai, Smale, Dotani, Hartmann and Heise, 1996; Shimura, 2000). Ebisawa et al. (2001) fitted a variety of spectral models to the spectrum obtained with the Solid State Spec-

Table 3.13: A summary of the X-ray properties of MR Vel, as obtained from published models fitted to the X-ray spectra. BB stands for blackbody and LTE refers to WD model atmosphere in local thermodynamic equilibrium, while non-LTE (NLTE) models allow departures from LTE. For more details regarding the particular WD atmosphere models, please refer to the relevant articles. The quoted values of the photospheric radius and bolometric luminosity are for a distance to the source of $d = 1$ kpc. However, the values corresponding to another distance can easily be determined using the expressions in the column headers. Where available, the χ^2 -value and degrees of freedom (dof) of the fit are also given. Where only a single value is given in the last column, it represents the reduced χ^2 -value.

Facility	Model	$\log g$ (cgs)	kT_{eff} (eV)	N_{H} (10^{20} cm^{-2})	R (d/kpc) (10^8 cm)	L_{bol} (d/kpc) 2 (10^{37} erg s $^{-1}$)	χ^2 /dof
<i>ROSAT</i> ^[1]	BB	–	43-55	140-190	~ 0.2	~ 20	18.4/17
<i>ROSAT</i> ^[2]	LTE	9	93	100	0.20	0.032	
<i>ROSAT</i> ^[2]	NLTE	9.25	72	100	0.40	0.050	0.78
<i>ASCA</i> ^[3]	BB	–	45	200	183	1800	605/75
<i>ASCA</i> ^[3]	BB ^{ab}	–	88_{-6}^{+10}	109 ± 11	$0.31_{-0.26}^{+0.51}$	$0.075_{-0.04}^{+0.013}$	50.8/69
<i>ASCA</i> ^[3]	LTE ^b	9	70	211	16	80	288/73
<i>ASCA</i> ^[3]	LTE ^b	10	86	165	1.8	2.3	288/73
<i>ASCA</i> ^[3]	LTE ^{bc}	10	117_{-2}^{+4}	71 ± 6	$0.036_{-0.019}^{+0.017}$	0.0031	79.2/71
<i>ASCA</i> ^[3]	NLTE ^b	9	74	87	0.16	0.01	393/73
<i>ASCA</i> ^[3]	NLTE ^b	10	92	85	0.088	0.007	343/73
<i>ASCA</i> ^[3]	NLTE ^{bc}	10	103 ^d	69_{-2}^{+6}	0.051	0.0038	51.7/73
<i>ASCA</i> ^[3]	LTE ^{be}	10	94	127	0.3		144/71
<i>BeppoSAX</i> ^[4]	NLTE	9.0	75.1 ± 0.6	82_{-8}^{+10}	$0.16_{-0.03}^{+0.05}$	$0.011_{-0.004}^{+0.007}$	90/15
<i>BeppoSAX</i> ^[4]	NLTE ^f	9.0	$75.4_{-0.3}^{+0.4}$	151_{-17}^{+18}	$0.1_{-0.3}^{+0.6}$	$0.41_{-0.22}^{+0.51}$	84/15
<i>BeppoSAX</i> ^[4]	NLTE ^f	9.5	91.6 ± 0.3	76 ± 4	0.098 ± 0.008	0.0079 ± 0.0014	165/15
<i>BeppoSAX</i> ^[4]	NLTE ^{f,g}	9.0	75.4 ± 0.3	142_{-15}^{+14}	$0.9_{-0.2}^{+0.3}$	$0.29_{-0.16}^{+0.18}$	47.6/14
<i>Chandra</i> ^[5]	BB	–	45	160	73	280	630/252
<i>Chandra</i> ^[5]	NLTE ^f	9	77.3	100	0.34	0.053	731/252
<i>Chandra</i> ^[5]	NLTE ^{f,h}	9	77.2	100	0.30	0.041	650/250

[1] Motch et al. (1994)

[2] Hartmann and Heise (1997)

[3] Ebisawa et al. (2001)

[4] Hartmann et al. (1999)

[5] Bearda et al. (2002)

^a With absorption edges at $1.01_{-0.02}^{+0.01}$ and $1.36_{-0.07}^{+0.09}$ keV.

^b With the oxygen and neon abundances in the interstellar absorption as free parameters.

^c With an additional absorption edge at ~ 1.02 keV.

^d 103 eV was the highest available temperature in this model.

^e Fitted simultaneously with a Raymond-Smith thin thermal plasma model with temperature 103 eV and emission measure $EM = n_e n_H V = 1.5 \times 10^{53} \text{ cm}^{-3}$, where n_e , n_H and V are the plasma electron density, hydrogen density and volume respectively.

^f Including the effect of line opacities.

^g Fitted simultaneously with a Collisional Ionization Equilibrium (CIE) model with temperature in the range from 1.2 keV down to 0.1 keV, corresponding to luminosities of $(0.0044-1.5) \times 10^{35} (d/\text{kpc})^2 \text{ erg s}^{-1}$ and emission measures of $EM = (0.15-60) \times 10^{56} \text{ cm}^{-3}$.

^h Fitted simultaneously with a CIE model with $kT = 175 \text{ eV}$ and $L = 4.9 \times 10^{34} (d/\text{kpc})^2 \text{ erg s}^{-1}$.

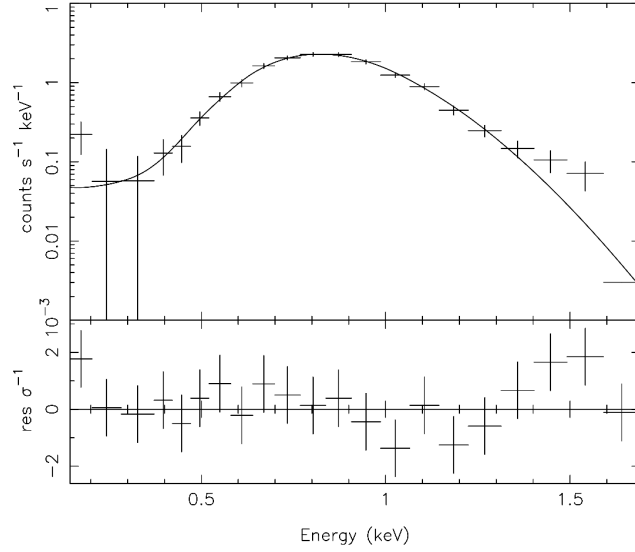


Fig. 3.45: The *ROSAT* PSPC spectrum of MR Vel on 20 November 1992, fitted with a NLTE WD model atmosphere. See Table 3.13 for the fit parameters. (Adopted from Hartmann and Heise, 1997, Fig. 6.)

trometer (SIS). Some of their fits are shown in Fig. 3.46, with the corresponding fit parameters in Table 3.13. They fitted a simple absorbed blackbody model, as well as a blackbody model with additional edge features, the latter fitting the spectrum quite well.

Ebisawa et al. (2001) also fitted several LTE and NLTE WD atmosphere models (without line opacities and line blanketing, only with continuum opacities) to the data, allowing the abundances of oxygen and neon in the interstellar absorption to be free parameters. For $\log g = 9$, they found a strong O VIII edge at 0.89 keV in the LTE model and a strong Ne IX edge at 1.2 keV in the NLTE model, none of which are present in the data. The data exhibit the Ne X edge at 1.36 keV instead. Hartmann et al. (1999) showed that in a more sophisticated NLTE model that includes metal line opacities, the Ne IX edge is weaker, but the O VIII edge stronger again, making it more similar to the $\log g = 9$ LTE model of Ebisawa et al.. Therefore the imperfect fit of the $\log g = 9$ models in Fig. 3.46 is unlikely to be caused only by a lack of sophistication of the models.

Consequently, Ebisawa et al. (2001) also considered LTE and NLTE models with $\log g = 10$, and found that the fit does not improve much, except if an absorption edge at 1.02 keV is added to the NLTE model. A two-temperature model was also applied, consisting of a LTE WD model atmosphere for the soft component, and a Raymond-Smith thin thermal plasma model for the hard component. Although this combined model accounted for the harder X-ray photons, the fit was not as good as the models containing the additional 1.02 keV edge. The apparent best fit model,

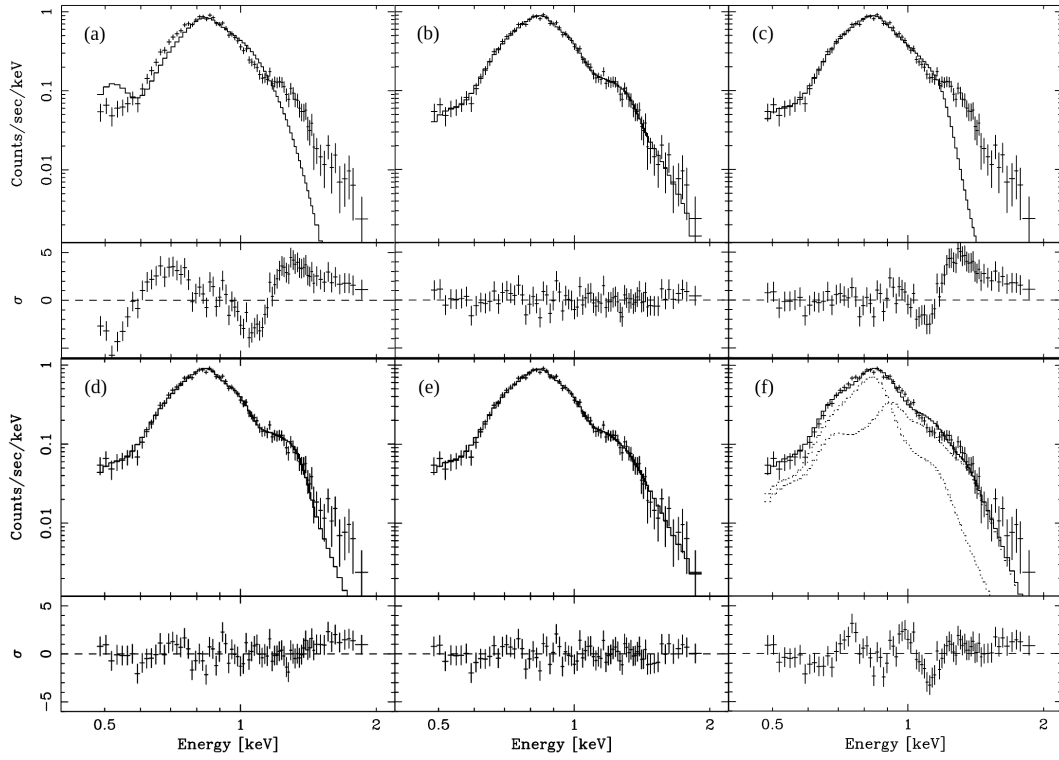


Fig. 3.46: The December 1994 *ASCA* spectrum of MR Vel, fitted with different models. (a) A simple absorbed blackbody. (b) An absorbed blackbody with an additional edges at 1.01 and 1.36 eV. (c) NLTE model with $\log g = 9$. (d) LTE model with $\log g = 10$ and an additional edge at 1.02 keV. (e) NLTE model with $\log g = 10$ and an additional edge at 1.02 keV. (f) LTE model with $\log g = 10$ (soft component) plus a Raymond-Smith plasma model (hard component). (Adopted from Ebisawa et al., 2001, Fig. 6, 7, 9, 10, 11 and 12.)

i.e. the NLTE model with a 1.02 keV edge, also indicates a very small radius and low luminosity at a distance of 1 kpc.

Ebisawa, Asai, Mukai, Smale, Dotani, Hartmann and Heise (1996) remarked on the fact that the minimum radius of a WD at the Chandrasekhar limit of $1.4 M_{\odot}$ is ~ 3000 km, and that the blackbody fit with additional edges requires a distance $\gtrsim 10$ kpc to be consistent with this minimum radius. The relatively high temperature required to fit the X-ray spectrum is also consistent with the WD being massive. However, it remains difficult to obtain a reasonably accurate estimate of the WD luminosity, as the bulk of the model atmosphere emission at soft X-ray energies is absorbed by the large column density.

Ebisawa et al. (2001) noted that, for their best fit NLTE model with the 1.02 keV edge, the luminosity even at $d = 10$ kpc is still at least an order of magnitude lower than expected for an SSS. They proposed that the WD is behind a Thomson thick cloud, almost fully ionized, which may represent matter previously ejected by the jet that is present in the system (see §3.3.2). X-ray photons scattered out of the line of sight by electrons in the cloud could significantly decrease the observed X-ray luminosity. The less ionized regions of the cloud might be responsible for the absorption edge at 1.02 keV.

Hoshi (1998) performed analytical calculations of the structure and associated luminosities and temperatures for various WD masses. Hoshi (1998) considered fit parameters $kT_{\text{eff}} \sim 98$ eV, $N_{\text{H}} \sim 0.9 \times 10^{22} \text{ cm}^{-2}$ and $L_{\text{bol}} \sim 1.6 \times 10^{35} (d/1 \text{ kpc})^2 \text{ erg s}^{-1}$ for a blackbody with additional absorption edges. It was consequently shown that the X-ray emission from MR Vel can be explained by an accreting WD with a mass of $\sim 1.4 M_{\odot}$ at a distance of $d \gtrsim 10$ kpc.

MR Vel was observed again with the *ROSAT* HRI for 55 ks over 4.1 d in June 1994 (Motch, 1996). Again, it is noted that the source does not seem to exhibit large-scale X-ray variability, with no X-ray count rate variations larger than $\sim 50\%$ since its discovery in 1991. The only significant variability observed in the June 1994 observation is a $\sim 40\%$ modulation on a time-scale of a few days, which may be related to the orbital period (see Fig. 3.51 and the discussion in §3.3.2). This source may therefore be one of the genuine steady burning SSSs according to the Van den Heuvel et al. (1992) model, or otherwise rising or declining on a time scale $\gtrsim 3$ yr.

MR Vel was observed with the Low and Medium Energy Concentrator Spectrometer (LECS and MECS) aboard *BeppoSAX* on 25-26 January 1997. Hartmann et al. (1999) fitted several NLTE WD model atmospheres to the data. They first fitted a $\log g = 9$ model with only continuum opacities, similar to the models used by

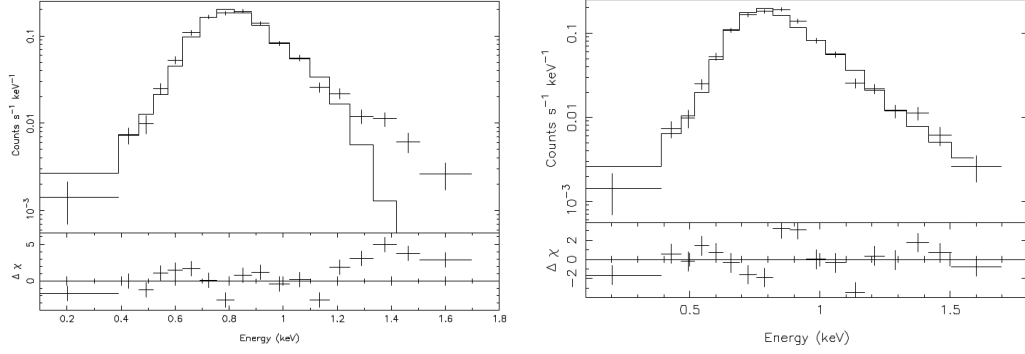


Fig. 3.47: The January 1997 *BeppoSAX* spectrum of MR Vel, fitted with different models. Left: A NLTE WD model atmosphere with line opacities included. Note that the model underestimates the data above 1.2 keV. Right: A better fit provided by adding a plasma component to the WD atmosphere. (Adopted from Hartmann et al., 1999, Fig. 2 and 7.)

Ebisawa, Asai, Mukai, Smale, Dotani, Hartmann and Heise (1996), and found similar source parameters to the ones derived from the *ROSAT* and *ASCA* spectra. They proceeded to calculate and fit $\log g = 9$ models with line opacities included to the *BeppoSAX* spectrum, although this did not drastically improve the fit (see Fig. 3.47). The main problem is that the data is in excess of the model above 1.2 keV. Hartmann et al. (1999) also considered high-gravity models with temperatures $\geq 10^6$ K, but these did not provide good fits to the data. All the relevant fit parameters of Hartmann et al. (1999) are listed in Table 3.13.

The final fit of Hartmann et al. (1999) statistically provided the best fit to the data. It consisted of a $\log g = 9.0$ NLTE fit (with line opacities), as well as a second spectral component: a Collisional Ionization Equilibrium (CIE) plasma with a temperature between 0.2 and 1.0 keV (Fig. 3.47). The addition of an optically thin plasma component to the optically thick atmospheric spectrum was also applied by Balman, Krautter and Ögelman (1998) in describing the X-ray spectrum of the nova V1974 Cyg. They argued that a fast wind collides with material that was ejected from the system at an earlier stage, providing the plasma component. The optical spectra of MR Vel do indicate the presence of a wind/jet in the system, providing justification for adding the CIE component.

The parameters of the NLTE model atmosphere did not change significantly upon the addition of the CIE component, as the former accounts for the bulk of the emission below 1.2 keV. The mass loss rate of a wind with terminal velocity v_∞ can be related to the plasma emission measure EM by (Balman et al., 1998)

$$\dot{m}_{\text{wind}} \approx 2v_\infty m_{\text{H}} \left(\frac{\text{EM}}{N_{\text{H}}} \right). \quad (3.3)$$

Hartmann et al. (1999) remarked that a simple approximation with

$\dot{m}_{\text{wind}} \sim 10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}$, $v_{\infty} \sim 1000 \text{ km s}^{-1}$ and $N_{\text{H}} \lesssim 10^{22} \text{ cm}^{-2}$ yield $\text{EM} \sim 10^{56} \text{ cm}^{-3}$, consistent with the emission measures derived from the *BeppoSAX* fits (and also the V1974 Cyg fits). One could therefore argue that the shape of the X-ray spectrum in itself provides tentative evidence for outflows in MR Vel.

Regarding the WD model atmosphere, Hartmann et al. (1999) concluded that NLTE models with effective temperatures above $\sim 10^6 \text{ K}$ ($\sim 86 \text{ eV}$) or below $\sim 7.5 \times 10^5 \text{ K}$ ($\sim 65 \text{ eV}$) can be excluded. If $\log g = 9$, the distance to the source is $\sim 4 \text{ kpc}$.

Bearda et al. (2002) presented the first high-resolution X-ray spectrum of MR Vel, obtained with the *Chandra* High Energy Transmission Grating Spectrometer (HETGS), utilizing the medium energy grating (MEG), on 14-15 November 2000. These authors fitted several models to the spectrum: (i) a blackbody, (ii) NLTE WD atmospheres similar to the models applied by Hartmann et al. (1999), and (iii) the NLTE component with a hot thermal plasma in collisional ionization equilibrium (see Table 3.13). However, none of these models provided a satisfactory fit to the spectrum, not even the model including the thin plasma component. Although the plasma component did add lines to the spectrum, the main problem was that no plasma component of any density or temperature succeeded in modelling the Fe XVII and Fe XVIII emission accurately.

The regions in the *Chandra* spectrum around the 5 strong, individually identifiable emission lines, as well as around the O VIII Lyman β line at $16.003 \text{ \AA}$, are shown in Fig. 3.48. These lines exhibit blueshifted absorption with velocities of $\sim 1500 \text{ km s}^{-1}$, indicating the presence of an outflow. It was noted that these velocities are lower than the jet velocities of $\sim 5000 \text{ km s}^{-1}$ derived from the optical spectra. Bearda et al. (2002) proposed that a better description of the X-ray spectrum might be obtained by combining an atmosphere spectrum with P Cyg profiles arising in a wind.

MR Vel was observed by *XMM-Newton* during two long pointings in late December 2000 (Motch, Bearda and Neiner, 2002), covering orbital phases of 0.82-0.98 and 0.31-0.47 according to the Schmidtke and Cowley (2001) ephemeris (which is discussed in §3.3.2). The X-ray lightcurve showed significant (full amplitude 10.8%) variability with a period of $5.11 \pm 0.22 \text{ h}$ (a similar time-scale as the optical variability, see §3.3.2), as well as variability on a shorter time-scale. However, Motch et al. (2002) did not find any periodicities in the 30-1000 s range.

The spectral analysis of the RGS data by Motch et al. (2002) confirmed the findings of Bearda et al. (2002): strong emission features in the spectrum, especially ionized Fe, could not be reproduced by the NLTE models that were used, and more sophisticated models including the effects of winds are required. The RGS spectrum with

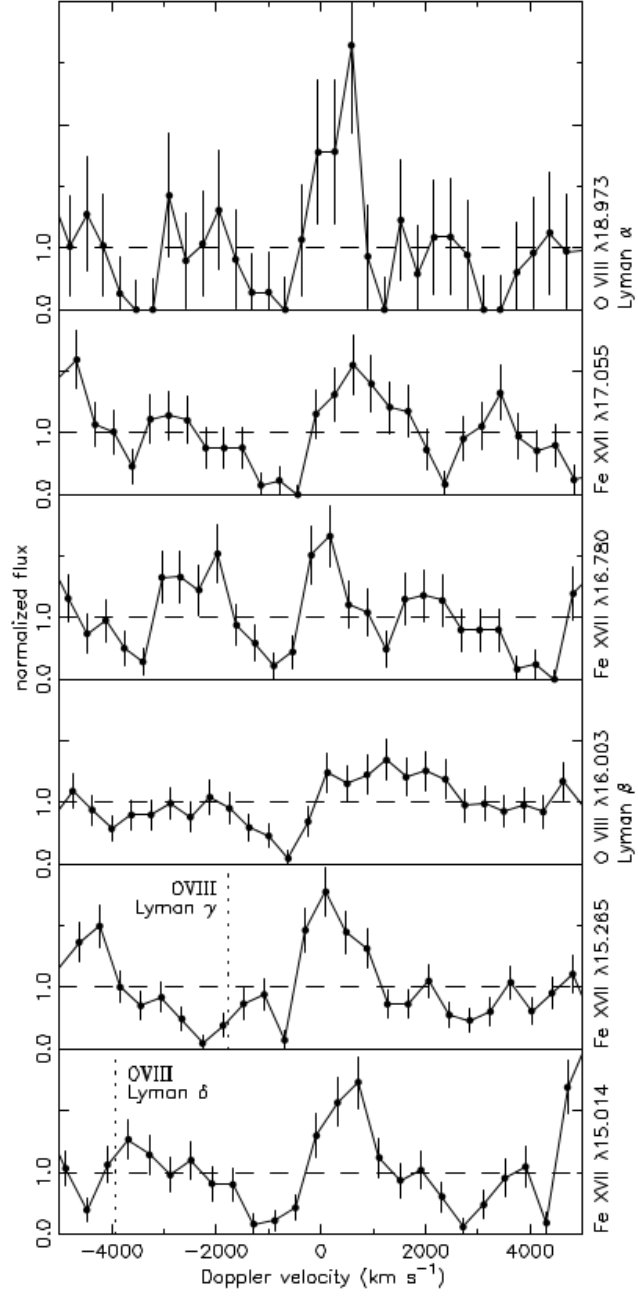


Fig. 3.48: Some prominent features in the *Chandra* spectrum of MR Vel in November 2000. P Cyg profiles with absorption velocities of $\sim 1500 \text{ km s}^{-1}$ are evident. (Adopted from Bearda et al., 2002, Fig. 2.)

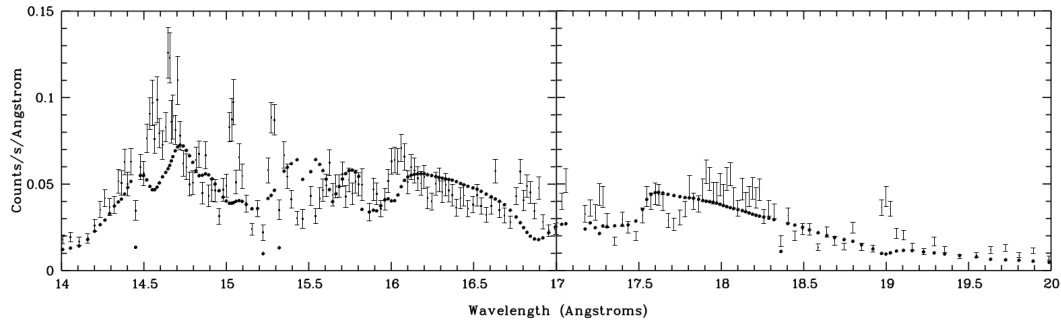


Fig. 3.49: The RGS1 spectrum of MR Vel recorded in December 2000 (small filled circles with error bars). The “best” fit NLTE model of Bearda et al. (2002) (fitted to *Chandra* data), but with $N_{\text{H}} = 1.28 \times 10^{22} \text{ cm}^{-2}$, is overplotted (larger filled circles). (Adopted from Motch et al., 2002, Fig. 5.)

the “best” fit NLTE model is shown in Fig. 3.49. Motch et al. (2002) confirmed the P Cyg profile in the O VIII Lyman α line at 18.97 Å, but with velocity of $\sim 2000 \text{ km s}^{-1}$ towards the blue, i.e. a slightly higher blueshift than derived by Bearda et al. (2002).

The X-ray flux in the second observation was $\sim 12 \pm 4\%$ higher over the whole energy range than in the first observation. Motch et al. (2002) deduced from the RGS spectra that the radial velocity shift of the emission lines between the first and second *XMM-Newton* pointings was $173 \pm 47 \text{ km s}^{-1}$. The orbital ephemeris of Schmidtke and Cowley (2001) predicts a difference of ~ 120 to 140 km s^{-1} between the He II $\lambda 4686$ radial velocities for observations at these orbital phases. This indicates that the He II $\lambda 4686$ line and the X-ray lines originate from approximately the same region in the system, i.e. that the He II emission may indeed represent the orbital motion of the WD. However, one must keep in mind the possibility that some He II emission may also originate from the heated side of the secondary or from the bulge of the accretion disc.

3.3.2 Optical

The ICRS coordinates of MR Vel are 09 25 46.00 –47 58 17.4. Motch et al. (1994) obtained optical spectra of this source at the ESO-MPI 2.2-m telescope from 17 to 25 April 1992 (see Fig. 3.50), and remarked that the X-ray and optical spectral properties of the source are very similar to that of the LMC supersoft source CAL 87. The source exhibited emission lines of H α , He II $\lambda 4686$ and also the Bowen blend (N III and C III) at 4640–4660 Å. The secondary star was not detected in the optical spectrum. The H α and He II $\lambda 4686$ lines exhibited narrow and broad components with radial velocity variations on time-scales of one day, with a full amplitude of $\sim 160 \text{ km s}^{-1}$. A systemic velocity of $+50 \pm 18 \text{ km s}^{-1}$ was estimated from the average radial velocity displacement of the He II $\lambda 4686$ line.

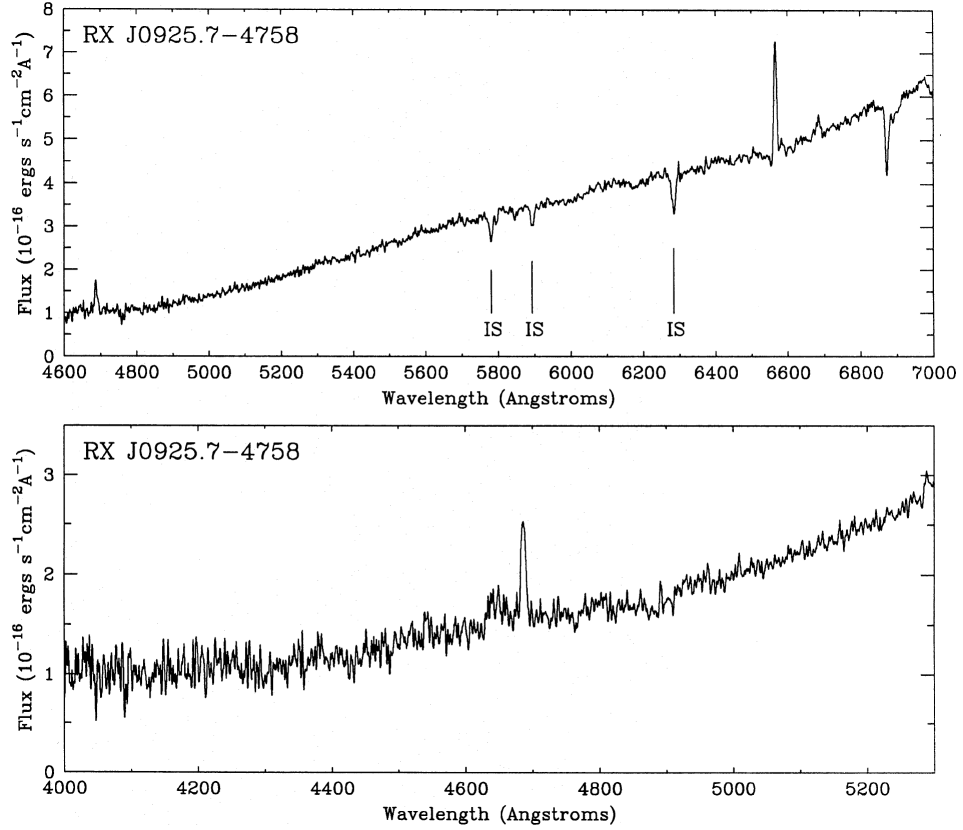


Fig. 3.50: The mean optical spectra of MR Vel recorded at the ESO-MPI 2.2-m telescope during April 1992. The dominant features in the low-resolution spectrum (top) are the $H\alpha$ and $\text{He II } \lambda 4686$ emission lines, and several interstellar (IS) absorption lines. In the medium-resolution spectrum (bottom), the $\text{He II } \lambda 4686$ emission line, as well as the Bowen complex at $\sim 4650 \text{ \AA}$. (Adopted from Motch et al., 1994, Fig. 5.)

The emission line properties are given in Table 3.14. The mean equivalent width of the Bowen complex is $-4.8 \pm 0.3 \text{ \AA}$. On some nights, the $H\alpha$ and $\text{He II } \lambda 4686$ emission seemed to be double-peaked, possibly due to emission from an accretion disc. The peak-to-peak velocity difference is 130 km s^{-1} and $170\text{--}300 \text{ km s}^{-1}$ for the two lines respectively. Although the intensity of the $H\alpha$ emission seems quite constant, the equivalent width of $\text{He II } \lambda 4686$ varies in the -4 to $-9 \text{ \AA}$ range.

Together with the spectroscopic measurements, Motch et al. (1994) obtained some photometric data with the 2.2-m, and also monitored the source on 21 and 23 November, and 15–20 December 1992 with the ESO Dutch 0.9-m telescope. During April and December, the V -magnitude varied in the 17.0–17.3 range on time-scales of several days, with the variability on time-scales of hours being $\lesssim 0.1 \text{ mag}$. The colour indices were $B - V = 2.01 \pm 0.06$, $V - R = 1.26 \pm 0.04$ and $V - I = 2.56 \pm 0.04$. The single measurements on 21 and 23 November yielded V -magnitudes of 17.11 and 17.06, and occurred respectively 1.1 d after the end of the pointed PSPC observation, and 0.26 d before the start of the pointed HRI observation. This detection

Table 3.14: The equivalent width (EW, in Å) and radial velocity (RV, in km s⁻¹) measurements of emission lines in MR Vel as measured at the ESO-MPI 2.2-m telescope in April 1992. The equivalent widths should be interpreted as negative values, as these are emission features. (Adopted from Motch et al., 1994, Table 1.)

Julian Day -2,448,700	EW He II λ 4686	EW H α	RV broad He II λ 4686	RV H α
31.5665	—	5.5 ± 0.3	—	+84
31.5921	5.7 ± 1.0	5.3 ± 0.1	+88	+86
32.5734	8.0 ± 3.0	5.5 ± 0.5	—	+166
32.6115	9.0 ± 1.5	4.5 ± 0.5	+126	+127
32.6350	4.5 ± 2.0	4.7 ± 0.3	+213	+156
35.5472	6.0 ± 1.0	—	-28	—
35.5743	8.0 ± 2.0	—	-63	—
35.6011	7.0 ± 1.0	—	-24	—
35.6435	—	5.0 ± 0.5	—	+8
36.5496	5.0 ± 1.0	—	+149	—
36.5894	5.4 ± 1.0	—	+104	—
36.6222	4.0 ± 1.0	—	+53	—
36.6632	6.0 ± 1.5	—	+129	—

of the source in its “normal” optical state quasi-simultaneously with an X-ray on-state suggested that the source might not be subject to the “limit cycle” behaviour characteristic of CAL 83 and RXJ0513, but such behaviour could not be excluded yet.

Motch et al. (1994) carried out a search for orbital modulations in their photometric data, and tentatively presented a value of ~ 3.5 d as the most probable orbital period. Folded on this period, the optical lightcurve exhibits a sharp maximum and a broad minimum. However, they could not correlate the observed radial velocity variations of the H α and He II λ 4686 well with this period.

By considering the slope of the reddened optical continuum (also comparing it to the continuum of RXJ0513), and by measuring the equivalent widths of the interstellar bands, Motch et al. (1994) derived a reddening of $E(B - V) \approx 2.0$, from which followed $N_{\text{H}} = (1.0\text{--}1.9) \times 10^{22} \text{ cm}^{-2}$. The dereddened magnitudes of MR Vel are $(B - V)_0 = -0.1$ and $(V - R)_0 = 0.06$ (Šimon, 2003).

MR Vel lies in the direction of the Vela Sheet molecular cloud (which is 425 pc away), and the source has a Galactic latitude a small distance above the denser regions of the CO emission (Motch et al., 1994). The total N_{H} towards this local cloud is $\approx 1.3 \times 10^{22} \text{ cm}^{-2}$, leading to the conclusion that MR Vel is located within or behind the Vela Sheet molecular cloud, but not too far away, yielding a lower limit to its distance (see Motch et al., 1994 and references therein). By assuming that the source does not radiate above the Eddington limit for a $1.4 M_{\odot}$ object, Motch et al. (1994) argued that the upper limit of the distance to the source is approximately

1-2 kpc.

Motch (1996) carried out new photometric and spectroscopic observations of the source on 5 consecutive nights in February 1994. By combining the newly obtained magnitudes with the data of 1992, he constrained the orbital period to 3.79 ± 0.24 d. Considering the Roche lobe geometry, the typical absolute magnitude of an SSS dominated by the irradiated accretion disc, and the absence of donor features in the spectrum, Motch (1996) proposed a donor star with an A0 to K4 giant or subgiant spectral type, with $M_V \sim 0$.

The H α and He II $\lambda 4686$ radial velocities are smoothly modulated with the orbital period, with semi-amplitudes of 63 ± 4 and 84 ± 30 km s $^{-1}$ respectively (Motch, 1996). The photometric maximum occurs when the line-emitting region is the closest to the observer. This is usually interpreted by a model where the emission lines are produced close to the primary, with the photometric modulation primarily caused by the changing aspect of the heated side of the donor. This is observed in the LMXBs Her X-1 and Sco X-1, with a similar phasing in the SSSs CAL 87 (Cowley et al., 1990) and RX J0019.8+2156 (Beuermann et al., 1995).

Motch (1996) also reported on additional V-band measurements obtained around the time of the June 1994 *ROSAT* observation (see Fig. 3.51). The time-scale of the X-ray variations seemed to be comparable to the photometric orbital period. Although this is not conclusive, it is therefore possible that the source exhibits orbital modulations in X-rays, similar to what is observed in the SMC 13 and GQ Mus (Kahabka, 1996). The optical maximum occurs approximately 0.1-0.2 cycles after the X-ray minimum. This is somewhat surprising, considering the possible “Her X-1 configuration”, but may be due to subtle changes in the X-ray spectral energy distribution along the line of sight.

No eclipses have so far been observed in the lightcurve of MR Vel. If the phasing of the events proposed above is correct, one can approximate the radial velocity semi-amplitude of the primary to be $K_1 \sim K_{\text{He II}} \approx 84$ km s $^{-1}$, i.e. a quite large mass function. Motch (1996) showed that all the non-eclipsing solutions with a WD primary imply a more massive secondary. For $M_{\text{wd}} = M_{\odot}$, and $i \geq 30^\circ$, the secondary mass is in the $1 M_{\odot}$ to $\geq 3.5 M_{\odot}$ range, with radius (5-8) $R_{\odot}$, consistent with the A0 to K4 giant or subgiant spectral type derived previously, and in good agreement with the Van den Heuvel et al. (1992) model.

MR Vel was observed at the ESO-Danish 1.54-m telescope for 1 hour on 7 June 1997, and for 2.8 hours on 8 June 1997 (Motch, 1998). The equivalent width of the H α emission line was -4.9 Å, and that of the He II $\lambda 4686$ line -7.8 Å. On 7 June,

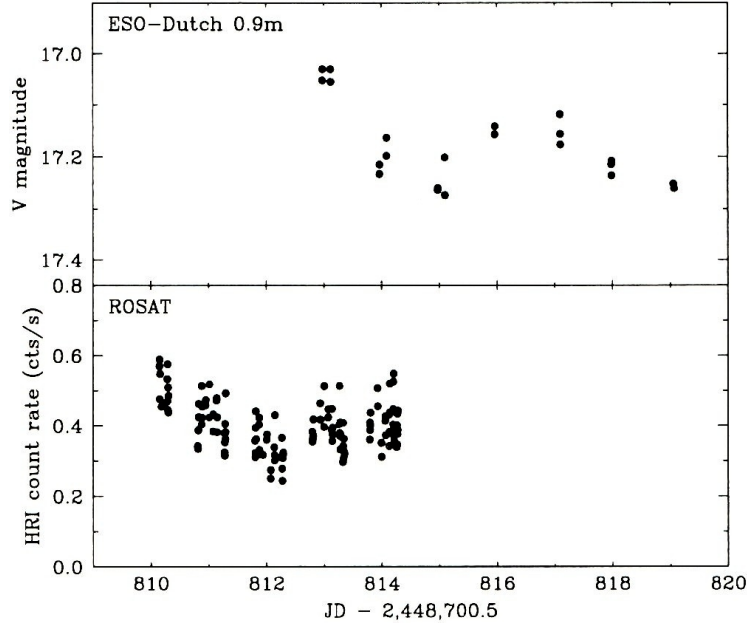


Fig. 3.51: The quasi-simultaneous optical and X-ray lightcurves of MR Vel in June 1994. (Adopted from Motch, 1996, Fig. 2.)

satellite features Doppler shifted by approximately 5200 km s^{-1} were observed in the $\text{H}\alpha$ emission (see Fig. 3.52), indicating the presence of bipolar jets in the system. No similar features were observed in the $\text{He II } \lambda 4686$ emission, with an upper limit of $< |-2.4| \text{ \AA}$ to the equivalent width of a potential similar blueshifted $\text{He II } \lambda 4686$ feature. On 8 June, the $\text{H}\alpha$ bipolar satellite features were no longer visible, with only faint emission at the position of the red feature, which was ascribed to $\text{He I } \lambda 6678$ and $\text{He II } \lambda 6683$. Therefore, they have only been detected on 1 night out of 23 on which the source have been observed spectroscopically since its discovery, while the jet in RXJ0513 is almost always visible.

The central $\text{H}\alpha$ line exhibited P Cyg profiles on both nights. It also exhibited variations in its “shoulders” on the first night, but not during the second. Southwell et al. (1997) remarked that precession of the jet in MR Vel would be possible for almost any WD mass, provided that the accretion rate is not higher than a few times $10^{-7} M_{\odot} \text{ yr}^{-1}$. Although the velocity shifts of the satellite features did not appear to be variable, the duration for which they could be observed was too short to exclude jet precession.

After correcting the red satellite feature on 7 June for the $\text{He I}/\text{He II}$ emission by subtracting the 8 June spectrum, it was found that the “pure” blueshifted and redshifted satellite features have comparable peak intensities (Motch, 1998). The equivalent width of the blue component was $-4.1 \text{ \AA}$. The V -magnitude of MR Vel was 17.133 ± 0.026 and 17.213 ± 0.016 on the two nights respectively, which is “nor-

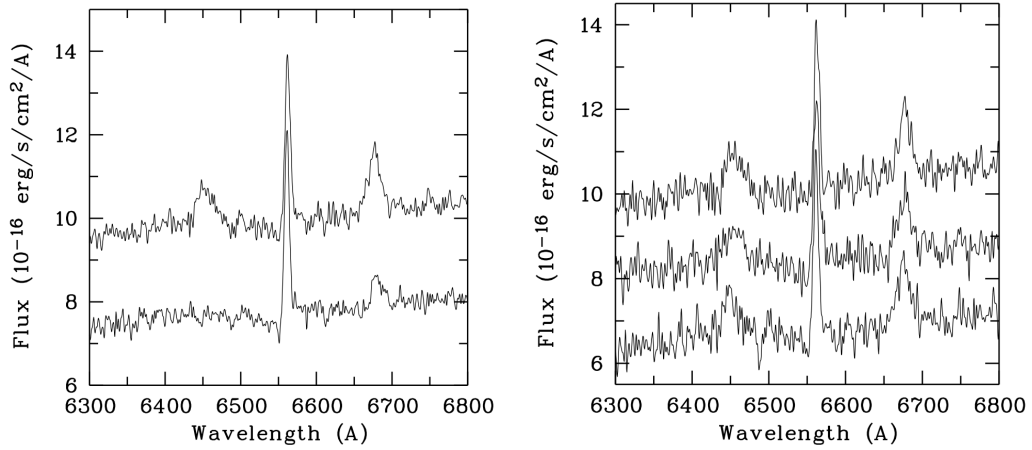


Fig. 3.52: Left: The mean spectra of MR Vel on 7 (top) and 8 June 1997 (bottom), showing the bipolar jet-associated satellite features in $H\alpha$ that were only observed on 7 June. The feature visible redward of $H\alpha$ on 8 June was ascribed to a blend of He I $\lambda 6678$ and He II $\lambda 6683$ emission, which is also blended with the red $H\alpha$ jet feature on 7 June. The 7 June spectrum has been shifted upwards by 3 flux units. Right: The 3 individual spectra on 7 June, shifted in flux for clarity. Time increases from bottom to top. (Adopted from Motch, 1998, Fig. 1 and 2.)

mal” for this source. Thus there is no evidence that the presence of the jet is in some way related to large-scale changes in the optical continuum emission.

Motch (1998) modelled the blue jet feature by considering a simple kinematic model. It was assumed that the jet is in the form of a cone with half-opening angle, and in the cone all atoms are moving at the same velocity, and uniformly per unit solid angle. He found that, at the 99% confidence level, the data are consistent with orbital inclinations in the 7-13 and 20-29° ranges, and half-opening angles in the 17-27 and 34-31° ranges. Alternatively, the velocity spread of the feature from 3800 to 5800 km s⁻¹ may be due to an intrinsic gradient in velocities in a much narrower jet, or the satellite features may be considerably broadened by Keplerian rotation of the material.

It was also pointed out by Motch (1998) that if the distance to MR Vel is ~ 10 kpc, the absolute magnitude is $M_V = -4$, much brighter than even RXJ0513. This could be due to the long orbital period, a larger secondary star, and the presumed large irradiated accretion disc in the system, as supported by the findings of Van Teeseling et al. (1997).

Optical spectra of MR Vel were obtained at the CTIO 4-m telescope on 4 November 1996 (Cowley et al., 1998). This spectrum is shown in Fig. 3.14 in §3.1.3, and the emission line properties are given in Table 3.15. The $H\alpha$ and He II $\lambda 4686$ emission lines and the Bowen blend are detected, but not the $H\alpha$ bipolar satellite features.

Table 3.15: The average equivalent widths of the MR Vel emission lines as observed with the CTIO 4-m telescope in November 1996 (Cowley et al., 1998, Table 5). The Balmer lines may include contributions from the He II Pickering lines.

Line ID	Measurement ($\text{\AA}$)
H α	-4.1
He II $\lambda 4686$	-5.3
He II $\lambda 4541$	-0.5
He II $\lambda 5411$	-0.5
O VI $\lambda 5290$	-0.8
FWHM of He II $\lambda 4686$	5.1

Table 3.16: Average properties of the spectral lines in MR Vel observed with the ESO 3.5-m NTT in 1998. The equivalent widths should be interpreted as being negative, as these are emission lines. (Adopted from Matsumoto and Mennickent, 2000, Table 1.)

Line	EW ($\text{\AA}$)	$FWHM$ (km s^{-1})
He II $\lambda 4686$	6-7	570-640
O VI $\lambda 5291$	≤ 0.5	~ 340
He II $\lambda 5413$	~ 1	~ 330
H α	4-5	370-410
He I $\lambda 6678$	~ 0.7	~ 450

H β is not observed in emission, and might even be weakly absorbing. Faint emission lines of O VI $\lambda 5290$, He II $\lambda 5411$ and C IV $\lambda\lambda 5801, 5812$ are also evident.

It is interesting to note that the Balmer and He II emission of MR Vel is much weaker than in other CBSSs like CAL 83 and RXJ0513, but the Bowen (N III, C III) and O VI $\lambda 5290$ emission strengths are comparable to that of the other sources. The ionized line strengths are probably related to the degree of ionization in the disc rather than actual element abundances. In particular, the O VI probably originates in the inner regions of the disc (see also Cowley et al., 1999).

Cowley et al. (1998) also obtained photometric measurements of MR Vel at the CTIO 0.9-m telescope. The V -magnitudes were found to be 17.19 ± 0.02 , 17.28 ± 0.02 and 17.31 ± 0.04 on 9, 11 and 12 December 1993 respectively.

Matsumoto and Mennickent (2000) reported on optical spectroscopic observations of MR Vel performed from 16 to 20 April, from 29-31 May 1998 and on 1 June 1998 with the ESO 3.5-m New Technology Telescope (NTT). Their spectra exhibited similar characteristics as published before, and they also noted emission features of He I $\lambda 6678$ and the Paschen series (Fig. 3.53). The emission line properties are listed in Table 3.16.

Radial velocities were measured for the H α and He II $\lambda 4686$, and fitted with sinusoidal curves. The radial velocity curves for these two lines were in phase. The

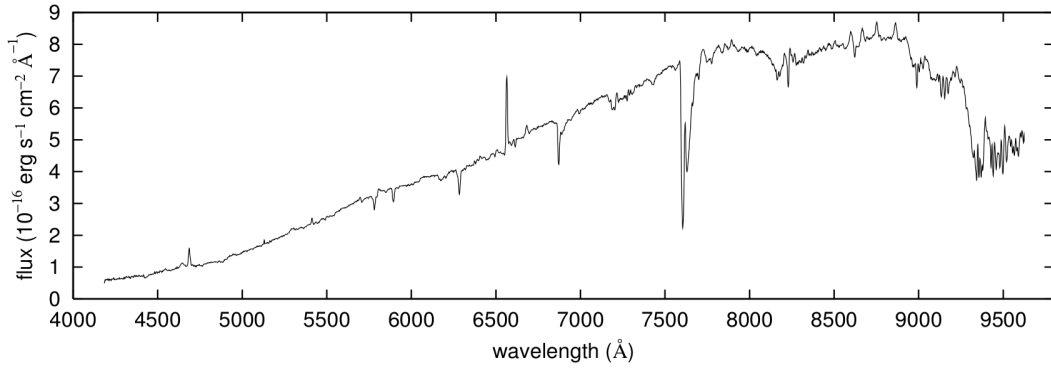


Fig. 3.53: The average optical spectrum of MR Vel observed in 1998 at the ESO 3.5-m NTT. (Adopted from Matsumoto and Mennickent, 2000, Fig. 1.)

Table 3.17: The result of sinusoidal fits to the radial velocity curves, showing the derived orbital period, semi-amplitude and systemic velocity (γ). (Adopted from Matsumoto and Mennickent, 2000, Table 3.)

Line	P_{orb} (d)	K_1 (km s $^{-1}$)	γ (km s $^{-1}$)
He II λ 4686	3.72 ± 0.11	122 ± 19	33 ± 10
H α	3.71 ± 0.09	70 ± 5	32 ± 3

results are shown in Table 3.17. The higher semi-amplitude of the He II λ 4686 line may be an indication that more highly ionized lines originate closer to the WD and might provide a better measure of the WD orbital motion.

Matsumoto and Mennickent (2000) adopted $M_{\text{wd}} \sim 1.44 M_{\odot}$, and found that the mass function based on the He II measurements indicate an evolved secondary with $M_2 \gtrsim 1.5 M_{\odot}$. Considering the absence in the optical spectrum of certain spectral lines that could be expected from the secondary, they constrained the properties of the secondary. They concluded that if $d \sim 7$ kpc, then $i < 65^\circ$, and the donor should be a giant of spectral type A3-F6. A later type secondary (around a G giant) is also plausible if $d \sim 4$ kpc.

Doppler tomography of the H α and He II λ 4686 lines showed that the dominant emission region for these lines is approximately between the two stars (Matsumoto and Mennickent, 2000). Also, the He II emission is markedly asymmetric. Thus, the radial velocity curves of the emission lines probably underestimate the real semi-amplitude of the WD orbital motion, and consequently provides only a lower limit for M_2 .

Schmidtke et al. (2000) obtained multi-colour optical photometry of MR Vel during five nights in February 1998 and eight nights in March 1999 with the CTIO 0.9-m telescope. By calculating periodograms of all their V-band data, and also the earlier data of Motch et al., and comparing it with the spectroscopic radial velocity curves,

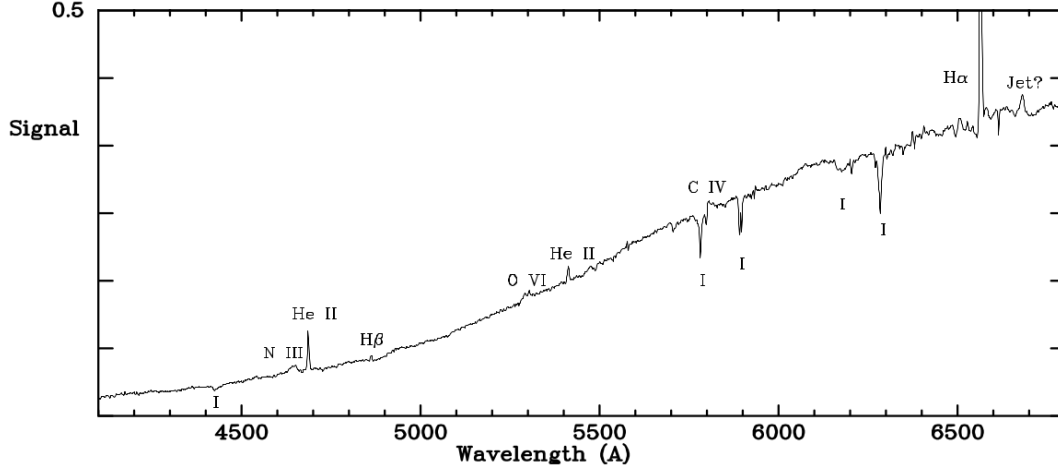


Fig. 3.54: Mean spectrum of MR Vel recorded in March 1999 at the CTIO 4-m telescope. I marks the interstellar absorption bands. (Adopted from Schmidtke et al., 2000, Fig. 4.)

an orbital period of 3.832903 ± 0.000003 d was derived. The full range of the orbital modulation is ~ 0.3 mag, and in none of the images of Schmidtke et al. (2000) was the source found to be brighter than $V = 17.05$.

Shorter time-scale variations of up to 0.1 mag in the photometric lightcurve were reported by Schmidtke et al. (2000). They also identified a quasi-periodic modulation with $P \sim 5.5$ h and an amplitude of 0.019 mag in the March 1999 data, and similar quasi-periods of ~ 6.2 and ~ 7.4 h with slightly smaller amplitudes in the February 1998 data. These short-term fluctuations did not seem to be dependent on the orbital phase. It was suggested that they may be related to luminous blobs moving through the accretion disc, or to changes in the accretion disc rim that occult a part of the inner region of the disc.

These authors determined colours of $V - R = 1.27$ and $B - V = 1.9$. Although the latter is poorly constrained, both colour indices are in good agreement with those reported by Motch et al. (1994). Within the error margins, the colours did not exhibit variations between observing runs or during the orbital cycle, or any correlation with the V -magnitude modulations.

Spectroscopic observations were obtained with the CTIO 4-m telescope quasi-simultaneously with the photometric observations in March 1999 (Schmidtke et al., 2000). The average spectrum is shown in Fig. 3.54. In addition to the same emission lines observed previously, $H\beta$ is now also observed in emission, superimposed on a broad absorption component. The feature at ~ 5290 Å, marked as O VI, seemed to be double peaked, but was proposed to be a blend of O VI $\lambda 5289$ and [Fe XIV] $\lambda 5303.6$. Average equivalent widths are given in Table 3.18.

Table 3.18: Average equivalent widths of the MR Vel emission lines as observed with the CTIO 4-m telescope in March 1999 (Schmidtke et al., 2000, Table 2). The Balmer lines may include contributions from the He II Pickering lines.

Line ID	Equivalent width (Å)
H α	−5.1
He II λ 4686	−5.7
He II λ 5411	−0.8
O VI λ 5290	−0.6

Table 3.19: The radial velocities semi-amplitudes and systemic velocities (V_0) of the emission lines of MR Vel as observed with the CTIO 4-m telescope in March 1999, obtained after folding on the photometric $P_{\text{orb}} = 3.832903$ d. (Adopted from Schmidtke et al., 2000, Table 3.)

Emission Line	K (km s $^{-1}$)	V_0 (km s $^{-1}$)	UT Date of Max. Velocity (1999 March)	Phase of Max. Velocity (photometric)	Mean Error of Fit (km s $^{-1}$)
He II λ 4686 (last obs. omitted).....	89.6 ± 3.2	43.1 ± 2.3	28.80 ± 0.02	0.85	12.4
He II λ 4686 (last two obs. omitted).....	88.9 ± 3.2	42.6 ± 2.3	28.80 ± 0.02	0.85	12.3
H α (all)	68.0 ± 1.8	42.2 ± 1.3	28.93 ± 0.02	0.89	7.3
H α (last two obs. omitted)	67.4 ± 1.9	41.8 ± 1.3	28.92 ± 0.02	0.88	7.3
O VI λ 5290	84.6 ± 27.6	5.4 ± 19.5	28.77 ± 0.02	0.85	102.2
“Jet,” 6680 Å	94.7 ± 10.4	...	28.92 ± 0.06	0.88	36.8

The radial velocities of the emission lines vary smoothly with the orbital period of 3.8 d, and have approximately the same phasing (see Table 3.19). The equivalent widths of H α and He II λ 4686 were also found to vary in phase with the V -magnitude, with maximum strength at orbital phase 0.5, when the primary is the closest to the observer. The amplitude of the He II λ 4686 variation was ~ 1 Å, and slightly smaller for H α , indicating that the H α line may originate further out in the disc, where it will suffer less occultation. Near maximum light, the H α line also becomes asymmetric, developing a red wing.

Schmidtke et al. (2000) found it unlikely that the feature at 6680.3 Å is He I λ 6678 or He II λ 6683, as no other He I lines are present in the spectrum, and its position, intensity and large width are not consistent with the other He II lines present. They therefore suggested that it might be the receding (red) H α satellite feature at a velocity of approximately +5350 km s $^{-1}$, while the approaching (blue) feature is not observed. They noted that the approaching satellite feature in RXJ0513 is also sometimes not seen in emission, and is at times even visible in absorption (Cowley et al., 1998). The appearance of the approaching component may thus depend on orbital phase. The observations of Schmidtke et al. were not at the same orbital phase as those of Motch (1998), but they note that their 6680 Å feature is the strongest at phases straddling the phase when the bipolar components were observed by Motch (1998). They argued that when Motch (1998) observed the 6680 Å feature, it was still the jet, but just at a weaker phase.

Utilizing the He II $\lambda 4686$ velocity to calculate the mass function, Schmidtke et al. (2000) arrived at a conclusion similar to that of previous authors: the orbital inclination should be $\lesssim 65^\circ$ and for a WD primary, the secondary is a giant with mass in the 1-1.6 $M_\odot$ range. The distance to the source should be ≥ 10 kpc. At very low inclinations, the donor mass and luminosity would be too high, making it visible in the optical spectrum. Schmidtke et al. (2000) further proposed that the photometric and radial velocity curves are modulated by an accretion disc rim, where the disc is thickened by the cooler (dark) mass transfer stream.

Determining the orbital period of MR Vel accurately had been complicated by the length of the period, which is comparable to the typical length of the observing runs dedicated to this source, and the uncertainty in the cycle counts when combining observations separated by months to years. Schmidtke and Cowley (2001) overcame this difficulty by obtaining photometric measurements on each clear night from 3 February to 2 July 2000 from the CTIO Yale 1-m telescope operated by the YALO (Yale, AURA, Lisbon, Ohio) consortium. By combining this extensive dataset with all available previous photometry, these authors derived the following orbital ephemeris for MR Vel:

$$T_{\min} = \text{HJD } 2451589.548 \pm 0.0008 + (4.0287822 \pm 0.0000026)E \text{ days} . \quad (3.4)$$

All the V -band data folded on this period is shown in Fig. 3.55. The folded lightcurve suggested a slightly asymmetric orbital modulation, with the decline somewhat more gradual than the rise. In addition to the orbital modulation, the previously reported ~ 0.1 mag variability on time-scales of several hours is also present in the YALO data.

Schmidtke and Cowley (2001) also re-evaluated the published $H\alpha$ radial velocities. A period analysis revealed a best fit of $P = 4.028716 \pm 0.000024$ d, which is very close to the photometric orbital period. The $H\alpha$ velocities folded on the updated photometric ephemeris are also shown in Fig. 3.55.

Motch et al. (2002) also reported on data from the Optical Monitor (OM) on *XMM-Newton* that were obtained simultaneously with the X-ray data (§3.3.1), as well as V - and B -filter photometry obtained at the ESO-Dutch telescope. During the *XMM-Newton* observations, it was found that, on average, $V = 17.23$ and $B - V = 1.88$, which is “normal” for this source. No correlation between the X-ray and optical variability could be deduced from the OM data, although it is noted that the signal-to-noise ratio (S/N) of the OM data was quite low.

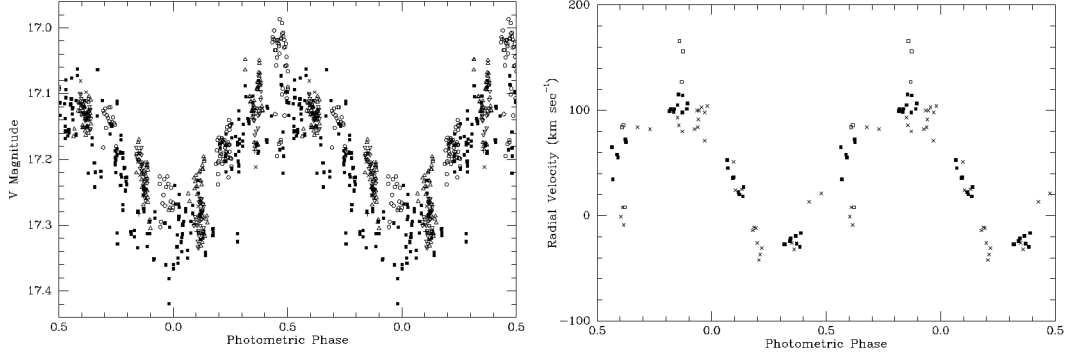


Fig. 3.55: Left: V -band photometry of MR Vel obtained from YALO, CTIO and ESO, folded on $P_{\text{orb}} = 4.0287822$ d. Right: $H\alpha$ radial velocities from the literature folded on the photometric P_{orb} . (Adopted from Schmidtke and Cowley, 2001, Fig. 2 and 4.)

3.3.3 Longer wavelengths

Infrared observations of MR Vel were obtained with the VLT-Antu on 20 May 2000 (Mennickent, Matsumoto and Diaz, 2003). These authors reported magnitudes of $J = 13.05$, $J - H = 0.76$ and $H - K = 0.53$. Using a colour excess of $E(B - V) = 1.94$, these correspond to dereddened colours of $J - H = 0.15$ and $H - K = 0.17$. These colours probably indicate the presence of a late-type companion or absorbing circumstellar material.

The infrared spectrum is presented in Fig. 3.56, and the spectral line measurements in Table 3.20. Mennickent et al. (2003) identified lines of neutral hydrogen (from the Paschen and Bracket series), He II and ionized oxygen. They noted the presence of P Cyg profiles in the Paschen β and Bracket γ lines. The blue edges of the absorption features are at velocities of -3000 km s^{-1} and -1500 km s^{-1} respectively relative to the main emission component, with the absorption minima at -755 km s^{-1} and -680 km s^{-1} for the two lines respectively. No spectral lines from the secondary star were observed.

Mennickent et al. (2003) calculated the expected spectral energy distribution of MR Vel based on a $1.4 M_{\odot}$ WD primary radiating as a blackbody, and accreting from an irradiated, supercritical accretion disc (see §2.4.5), including irradiation of the companion star. They considered the parameter δ , representing the “opening angle” of the disc, with small values of δ corresponding to a standard thin disc. According to Hanamoto, Ioroi and Fukue (2001), $\delta > 0.3^{\circ}$ for accretion rates $\geq 10^{-7} M_{\odot} \text{ yr}^{-1}$. Mennickent et al. (2003) found that, for an inclination angle in the $45\text{--}65^{\circ}$ range, the spectral energy distribution of the optical and infrared data indicate a distance to the source between 2 and 4 kpc for a wide range of δ (0.3–5). Even for an unrealistically high value of $\delta = 20$, the distance is $d \sim 5\text{--}6$ kpc. This would yield an absolute magnitude for the source between $M_V = -2$ and 0.

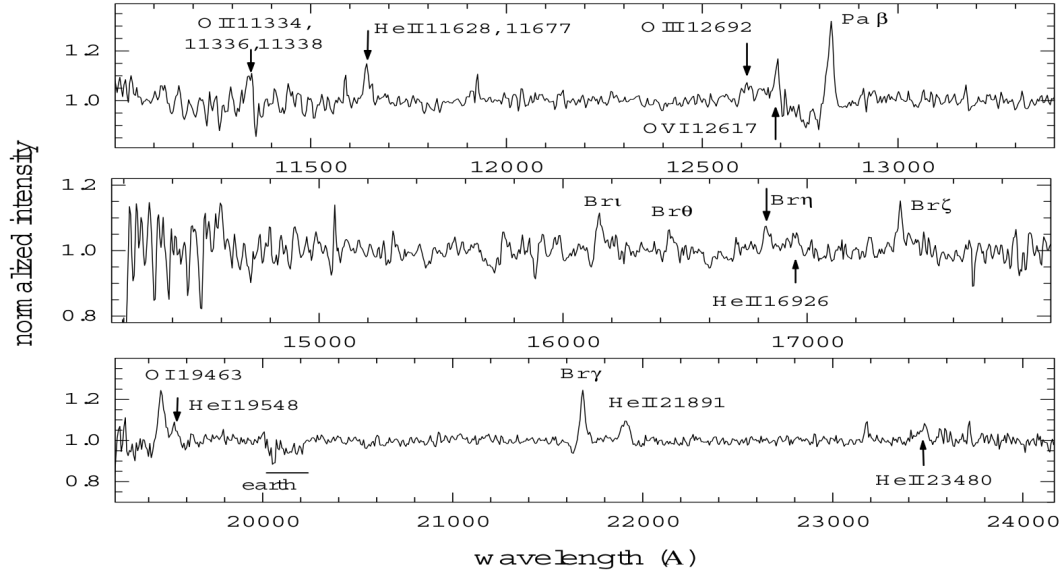


Fig. 3.56: The *J*-, *K*- and *H*-band (top to bottom) continuum normalized infrared spectra of MR Vel obtained on 20 May 2000 at the VLT-Antu. (Adopted from Mennickent et al., 2003, Fig. 2.)

Table 3.20: Infrared spectral line properties of MR Vel as recorded on 20 May 2000 at the VLT-Antu. The quoted peak intensity is in units of $10^{-15} \text{ erg s}^{-1} \text{ cm}^{-2} \text{ Å}^{-1}$. (Adopted from Mennickent et al., 2003, Table 1.)

Line	Intensity	EW (Å)	$FWHM$ (km s $^{-1}$)
O II 11334-38	2.2	−2	N/A
He II 11628-77	2.3	−2	N/A
O VI 12617	2.0	N/A	N/A
O III 12692	2.2	N/A	N/A
Paβ	2.4	−5	367
Brγ	1.7	−3	555
Brθ	1.6	−2	627
Brη	1.5	−2	560
He II 16926	1.5	−1	540
Brζ	1.5	−4	440
O I 19463	1.3	−9	564
He II 19548	1.2	−3	520
Brγ	1.0	−9	473
He II 21891	0.9	−4	740
He II 23480	0.7	N/A	N/A

From the extensive literature review provided in this chapter, it is evident that CAL 83, RXJ0513 and MR Vel share various properties. The first and obvious similarity is the supersoft nature of their X-ray spectra, which is produced by nuclear burning in a shell on the surface of a WD accreting at a very high rate. All three sources also exhibit emission lines expected to originate mainly in an accretion disc. The optical spectra present evidence of outflows in the form of P Cyg profiles, and also emission features Doppler shifted by several thousand km s^{-1} . These velocities are comparable to the escape velocity of a WD, implying that our line of sight is more or less aligned with the outflow vector. Together with other observational properties, this indicates relatively low orbital inclinations. P Cyg profiles were also reported in the X-ray spectra of RXJ0513 and MR Vel. CAL 83 and RXJ0513 exhibit anti-correlated modulations in the long-term X-ray and optical light curves, which are ascribed to effective temperature changes associated with the contraction and expansion of an extended WD photosphere. Such cycles have not yet been observed in MR Vel. If they do exist, it will be on a much longer time-scale than in the other two sources, and will only be revealed by ongoing X-ray and optical monitoring.

The rapid mass transfer in these supersoft X-ray binaries provides excellent environments in which to study the accretion phenomena discussed in Chapter 2, like MHD turbulence, accretion winds and collimated outflows. The high rate of mass transfer on the thermal time-scale of the donor (millions of years) creates the possibility that the WDs in these systems might have been spun up significantly to short spin periods by accretion disc torques. The energetic environment on the WD surface may also increase the probability of triggering unstable oscillations in the envelope or even non-radial pulsations of the WD itself, that may manifest as periodicities in the X-ray lightcurves.

Consequently, the observational results obtained from X-ray variability studies of CAL 83 (Chapter 4), and optical spectroscopy and photometry of CAL 83, RXJ0513 and MR Vel will be presented (Chapters 5 and 6).

Chapter 4

XMM-Newton observations of CAL 83

CAL 83 in the Large Magellanic Cloud has been observed extensively by *XMM-Newton*. This satellite observatory hosts a total of six science instruments, i.e. the three European Photon Imaging Cameras (EPIC), two Reflection Grating Spectrometers (RGS) and the Optical Monitor (OM), and they can all be operated at the same time. CAL 83 is known to exhibit X-ray, UV and optical variability on various time-scales, making *XMM-Newton* the ideal facility for simultaneous studies of the X-ray to near-UV/optical properties of the source, especially their variability.

This chapter will be subdivided in seven sections. In §4.1, the key properties of the *XMM-Newton* instruments will be highlighted, followed by a summary of the existing archival observations of CAL 83 in §4.2. A detailed analysis of the *XMM-Newton* data will be presented in the subsequent four sections, each containing the methodology and results pertaining to one focus area in the investigation. In §4.3, the X-ray and optical lightcurves of the individual observations will be introduced, and the correlation between them investigated. The discovery and characteristics of a ~ 67 s X-ray periodicity will be presented in §4.4, and the presence of various X-ray periodicities on slightly longer time-scales of several minutes to hours will be discussed in §4.5. In §4.6, the *XMM-Newton* measurements will be correlated with the *I*-magnitude lightcurves obtained as part of the OGLE (Optical Gravitational Lensing Experiment) Project, in order to put these measurements into context with the long-term variability of the source. Finally, an integrated discussion of all the observational results will be presented in §4.7.

4.1 The X-ray Multi-Mirror Mission (*XMM-Newton*)

XMM-Newton was launched in 1999 and is operated by the European Space Agency (ESA). It is in an elliptical high Earth orbit, and this enables uninterrupted observations of up to ~ 40 hours, making it ideal for studying source variability. The payload consists of 3 Wolter type-1 X-ray telescopes and also a co-aligned 30 cm op-

tical/UV telescope to enable simultaneous X-ray and optical/UV observations. Two of the X-ray mirror modules are equipped with Reflection Grating Arrays, while the third module has an “open” light path. There are a total of six science instruments on board *XMM-Newton*, i.e. the three European Photon Imaging Cameras (EPIC), two Reflection Grating Spectrometers (RGS) and the Optical Monitor (OM), and they can all be operated simultaneously. Schematic diagrams of the optical paths are shown in Fig. 4.1. A brief overview is given below, compiled from the *XMM-Newton* Users Handbook (XMM-Newton Community Support Team and Science Operations Center Team, 2013), with a summary of the technical characteristics of the instruments provided in Table 4.1. For further information, the reader is referred to the Users Handbook, as well as the instrument papers cited below.

4.1.1 The European Photon Imaging Cameras (EPIC)

These three cameras do not have gratings in their optical paths, but because of the intrinsic energy resolution of the CCDs, simultaneous X-ray imaging and non-dispersive X-ray spectroscopy can be achieved. As the CCDs are also sensitive to infrared, visible and ultraviolet light, each of the EPIC cameras has three separate optical blocking filters available, i.e. thick, medium and thin, the choice of which will depend on the optical magnitude of the target. The EPIC cameras function in photon counting mode, i.e. an event list is produced with the detection time, energy, position and other properties of each detected photon event.

EPIC pn

The EPIC pn detector is mounted in the focal plane of the mirror module that does not contain a grating. It consists of an array of 12 back-illuminated CCDs (Strüder et al., 2001). It has the highest effective area of the X-ray detectors, and is the one best suited for timing observations.

EPIC MOS

In each of the two mirror modules containing the grating arrays, 40% of the radiation is dispersed by the grating array onto a linear strip of CCDs (the RGS detectors, see the next section), while 44% of the radiation is directed, non-dispersed, at an EPIC MOS (Metal Oxide Semi-conductor) array mounted at the prime focus. There is one MOS array for each of these modules, i.e. MOS1 and MOS2. Each MOS array consists of 7 front-illuminated CCDs (Turner et al., 2001).

4.1.2 The Reflection Grating Spectrometers (RGS)

The light dispersed through the two Reflection Grating Assemblies (RGAs) in the two respective mirror modules is directed to the RGS Focal Cameras (RFCs). In

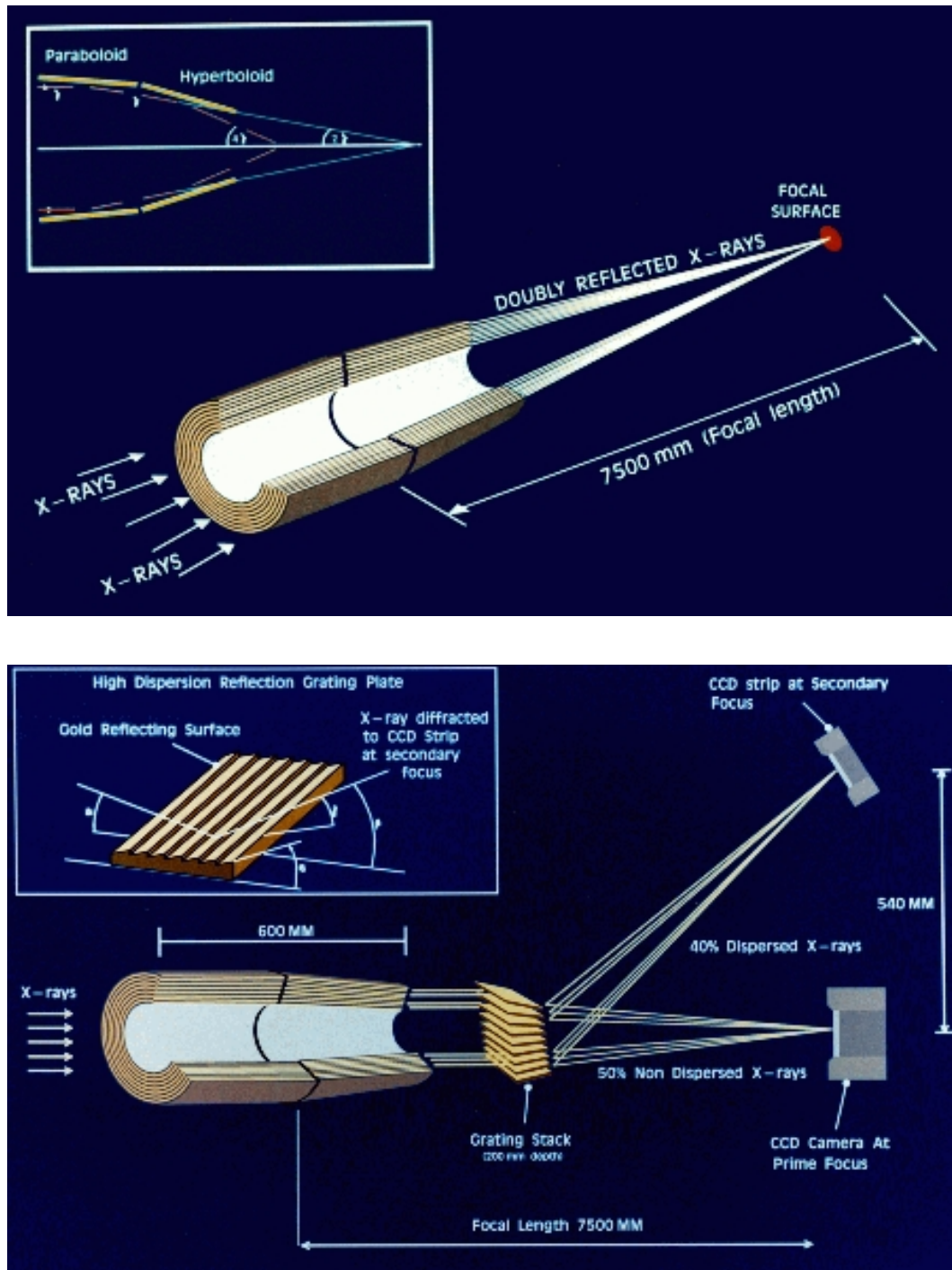


Fig. 4.1: Top: The light path in the *XMM-Newton* X-ray telescope without a grating assembly, with the EPIC pn camera in focus. Bottom: An illustration of the nature of the light path in the two X-ray telescopes containing grating assemblies. Of the incident light, 44% is transmitted to the MOS camera at the primary focus, and 40% reflected to the RGS focal cameras. (Adopted from the *XMM-Newton* Users Handbook, 2013, Figures 2 and 3 , pp. 9-10.)

Table 4.1: A summary of the properties of the *XMM-Newton* instruments. (Table entries and selected notes adopted from the *XMM-Newton* Users Handbook, 2013, Table 1, p. 8.)

Instrument	EPIC MOS	EPIC pn	RGS	OM
Bandpass	0.15–12 keV	0.15–12 keV	0.35–2.5 keV ⁽¹⁾	180–600 nm
Orbital target vis. ⁽²⁾	5–135 ks	5–135 ks	5–135 ks	5–145 ks
Sensitivity ⁽³⁾	$\sim 10^{-14}$ ⁽⁴⁾	$\sim 10^{-14}$ ⁽⁴⁾	$\sim 8 \times 10^{-5}$ ⁽⁵⁾	20.7 mag ⁽⁶⁾
Field of view (FOV)	30' ⁽⁷⁾	30' ⁽⁷⁾	$\sim 5'$	17'
PSF (<i>FWHM/HEW</i>) ⁽⁸⁾	5" / 14"	6" / 15"	N/A	1.4"–2.0"
Pixel size	40 μm (1.1")	150 μm (4.1")	81 μm (9×10^{-3} Å) ⁽⁹⁾	0.476513" ⁽¹⁰⁾
Timing resolution ⁽¹¹⁾	1.75 ms	0.03 ms	0.6 s	0.5 s
Spectral resolution ⁽¹²⁾	~ 70 eV	~ 80 eV	0.04/0.025 Å ⁽¹³⁾	350 ⁽¹⁴⁾

1) In the $-1.$ grating order (wavelength range: 5–35 Å; λ [Å] $\times$ E [keV] = 12.3984).

2) Total time available for science per orbit; minimum of 5 ks (excl. overheads)

in order to ensure observatory efficiency. *XMM-Newton* science observations can only be performed outside the Earth's radiation belts.

3) After 10 ks.

4) In the range 0.15–12.0 keV, in units of $\text{erg s}^{-1} \text{cm}^{-2}$.

5) O VII 0.57 keV line flux in photons $\text{cm}^{-2} \text{s}^{-1}$, for an integration time of 10 ks and a background of 10^{-4} photons $\text{cm}^{-2} \text{s}^{-1} \text{keV}^{-1}$.

6) $5\text{-}\sigma$ detection of an A0 star in 1000 s.

9) In spectroscopy mode (standard 3×3 pixel on-chip binning applied).

10) 1" with 2×2 binning in default configuration mode.

11) In fast data acquisition mode (i.e., fast mode for OM and timing mode for EPIC, spectroscopy mode for RGS1 [1.2 s for RGS2], reading out only one of nine CCDs). The EPIC pn burst mode offers an even higher timing resolution of 7 μs , but has a very low duty cycle of 3%.

12) At 1 keV energy. At the energy of Fe K α (6.4 keV), the energy resolution of both EPIC cameras is ca. 150 eV.

13) In $-1.$ and $-2.$ order, resp.; at 1 keV, this corresponds to 3.2/2.0 eV (*HEW*).

14) Resolving power ($\lambda/\Delta\lambda$) with UV and optical grism.

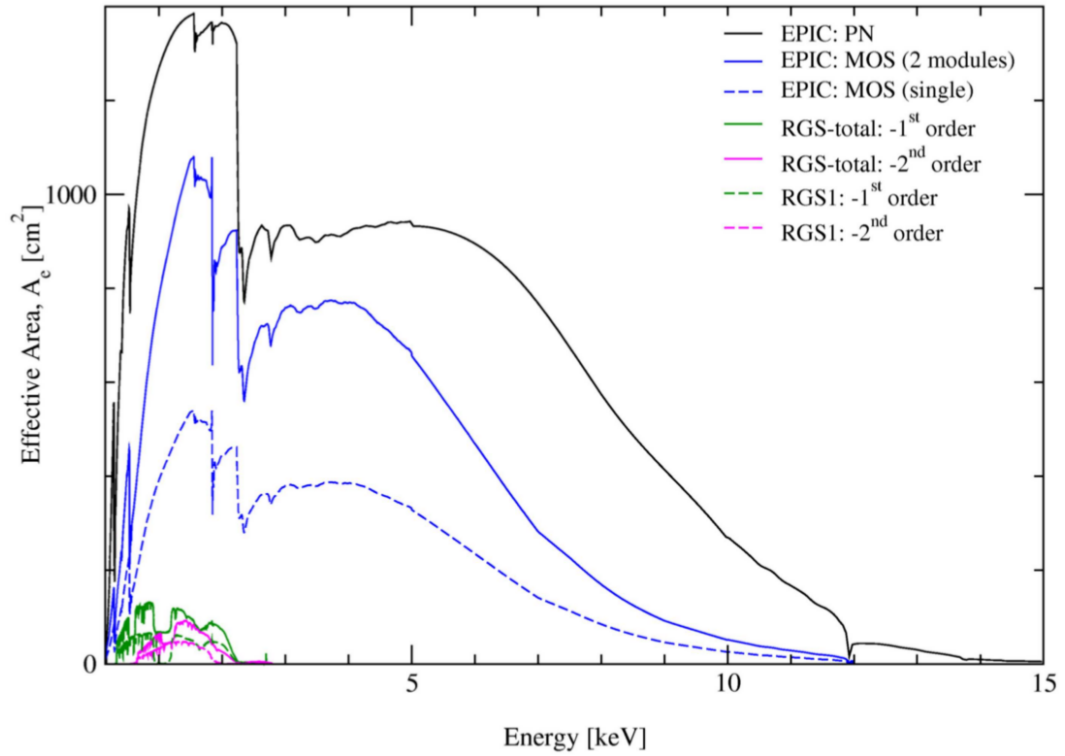


Fig. 4.2: The net effective area of all the X-ray telescopes on *XMM-Newton*. (Adopted from the *XMM-Newton* Users Handbook 2013, Figure 11, p. 16.)

each mirror module, the RFC consists of a linear array of 9 MOS CCDs similar to the EPIC MOS CCDs, but back-illuminated to maximise the soft energy response. The RGS is best suited to high resolution X-ray spectroscopy in the soft energy range, i.e. 0.33-2.5 keV (den Herder et al., 2001). The RGS events can also be used for timing purposes, albeit with lower timing resolution than the EPIC cameras.

4.1.3 The Optical Monitor (OM)

The OM has a mirror diameter of 30 cm, and its sensitivity range is 180-600 nm (Mason et al., 2001). It allows simultaneous optical/UV and X-ray observations. It has a filter wheel containing 10 elements, of which only one can be in the light path at a given time. There are 7 filters, 2 grisms and 1 blocked position. The available filters are white, UVW2, UVM2, UVW1, U, B and V, and their throughputs are shown in Fig. 4.3. The optical grism has a dispersion of ~ 5 Å per pixel, while the dispersion of the ultraviolet grism is ~ 2.5 Å per pixel. Their inverse sensitivity functions are shown in Fig. 4.4.

The OM can be operated in two modes. In Imaging Mode, spatial coverage is emphasized, but timing information on the individual photons is not stored. However, in Fast Mode, an accumulated 2-dimensional image is not stored, but an event list for a small window centred on the target is created, similar to the event files of the

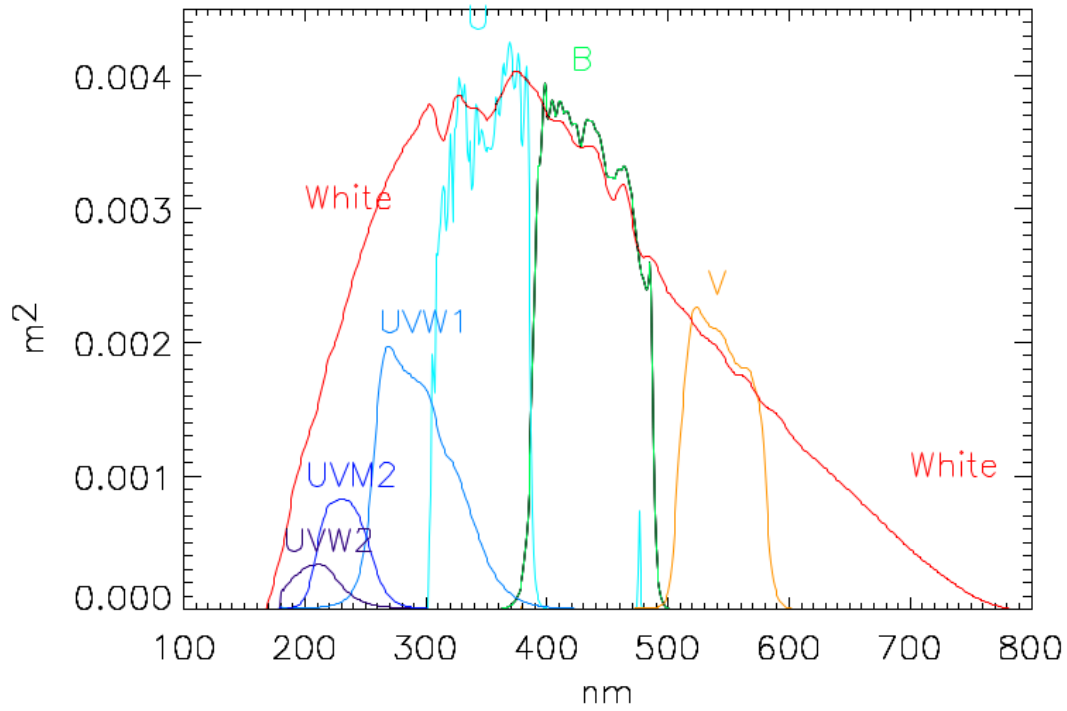


Fig. 4.3: The throughput curves of the 7 OM filters in terms of effective area (in m^2), folded with the sensitivity of the detector. (Adopted from the *XMM-Newton* Users Handbook, 2013, Figure 103, p. 135.)

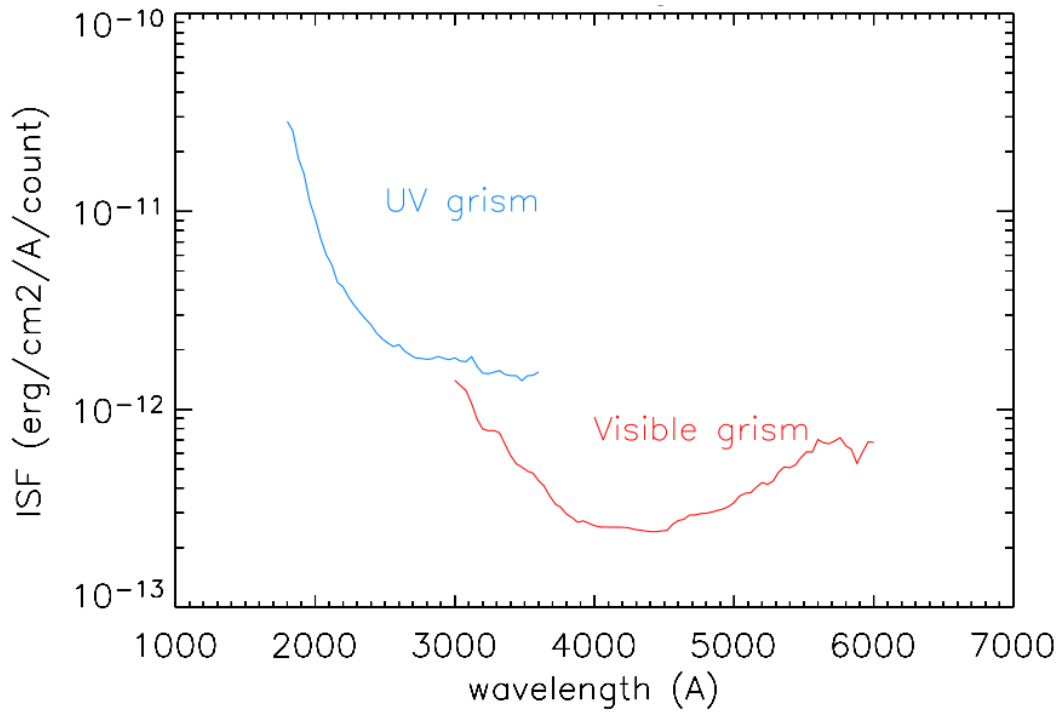


Fig. 4.4: The inverse sensitivity functions of the OM grisms. (Adopted from the *XMM-Newton* Users Handbook, 2013, Figure 104, p. 134.)

X-ray detectors (i.e. photon counting mode). The latter mode is very useful for observing rapidly variable sources. The two modes are often operated simultaneously, with Imaging Mode windows as well as a small Fast Mode window.

4.1.4 The *XMM-Newton* time system

When considering the arrival times of photons from an astronomical object, it is important to know (1) the “reference frame” or geometric location from which the time is measured, and (2) the “time standard”, which represents the way in which a specific clock ticks, as well as its defined zero point. The time stamp of the photon event is then a combination of these two factors.

The time system used by *XMM-Newton* is Terrestrial Time (TT), although the Coordinated Universal Time (UTC) of an observation is often also given in the FITS file header. The time stamps of the events themselves in the calibrated event files correspond to Mission Elapsed Time (MET) in seconds (in the local reference frame) since the reference time 1998-01-01 00:00:00.000 (TT) with the geocentre as reference point. Expressing this as a Julian Date yields $JD_{TT}(REF)$ 2450814.5 (*XMM-Newton* Users Handbook, 2013, p. 172).

However, the most practical absolute time system for recording astrophysical events is the Barycentric Julian Date (BJD) in the Barycentric Dynamical Time standard (TDB). The general method with which to convert a photon arrival time in terms of JD in UTC to a BJD in TDB is given by e.g. Eastman, Siverd and Gaudi (2010):

$$BJD_{TDB} = JD_{UTC} + \Delta_{R\odot} + \Delta_C + \Delta_{S\odot} + \Delta_{E\odot} , \quad (4.1)$$

where $\Delta_{R\odot}$, Δ_C , $\Delta_{S\odot}$ and $\Delta_{E\odot}$ represent the Rømer delay, clock correction, Shapiro delay and the additional Einstein delay respectively. Each of these terms will now be discussed separately.

Because the speed of light is finite, the time of arrival of a stellar photon at the Earth will depend on the position of the Earth in its orbit around the Sun, and can be advanced or delayed by 8.3 minutes. Therefore, the arrival time should rather be referred to the position of the solar system barycentre (SSB). This change of reference frame to the SSB is accomplished by adding the Rømer delay ($\Delta_{R\odot}$), which represents the arrival time difference between a photon travelling to an earthly observer (or satellite), and a photon travelling directly to the SSB.

The third term on the right, Δ_C , is the clock correction and enables the conversion

from one time standard to another. From UTC to TDB, the correction will be

$$\Delta_C = N + 32.184 \text{ s} + (\text{TDB} - \text{TT}) . \quad (4.2)$$

Here N represents the current amount of leap seconds, and converts UTC to TAI (International Atomic Time). The second term on the right is a simple offset used to convert TAI to TT, and the last term obviously converts TT to TDB. (Evidently, the terms used will depend on the required conversion; e.g. converting from TT to TDB will involve only the last term.) The $\text{TDB} - \text{TT}$ term performs a correction to TT for the delay because of time dilation and gravitational redshift from the movement of the Sun and the other bodies in our solar system (known as the Einstein delay). It is a predominantly periodic correction with magnitude of a few milliseconds at most and can be approximated by

$$\text{TDB} - \text{TT} = 0.001658 \sin g + 0.000014 \sin 2g \text{ seconds} , \quad (4.3)$$

where (e.g. Rots, Boyd and Markwardt, 2010)

$$g = 357.53^\circ + 0.9856003^\circ(\text{JD}_{\text{TT}} - 2451545.0) . \quad (4.4)$$

In Eq. (4.1), the Shapiro delay ($\Delta_{S\odot}$) compensates for the delay experienced by light passing near a massive object (the Sun in this case) according to general relativity. The Einstein delay correction for relativistic effects for an observer at the geocentre is already incorporated in the clock correction to TDB described above. The $\Delta_{E\odot}$ term in Eq. (4.1) represents an additional term to the Einstein delay that is necessary when the observing system is not at the geocentre, but at a certain position on the surface of the Earth or on a satellite.

Converting the *XMM-Newton* reference time to a time in BJD_{TDB} is simply achieved by applying the relevant clock correction:

$$\begin{aligned} \text{BJD}_{\text{TDB}}(\text{REF}) &= \text{JD}_{\text{TT}}(\text{REF}) + (\text{TDB} - \text{TT}) \\ &= 2450814.5 - 0.00000000083 \\ &= 2450814.49999999917 , \end{aligned} \quad (4.5)$$

where $\text{TDB} - \text{TT}$ was obtained from Eq.'s (4.3) and (4.4). It is evident that the difference between $\text{BJD}_{\text{TDB}}(\text{REF})$ and $\text{JD}_{\text{TT}}(\text{REF})$ is negligible for most practical purposes.

The software used for the processing of *XMM-Newton* data is the Science Analysis

System (SAS). More information about SAS can be found on the SAS website¹, and in the Users Guide (XMM-Newton Science Operations Centre Team, 2014). By performing the necessary corrections shown in Eq. (4.1), the SAS task **barycen** converts the recorded MET_{TT} of an *XMM-Newton* event in the local reference frame to an MET referred to the SSB (MET_{TDB}), by utilizing the right ascension and declination of the target, as well as the information pertaining to the satellite orbit at that moment in time. Throughout this chapter, wherever the full BJD_{TDB} of a certain event is given, it was obtained by adding the MET_{TDB} to BJD_{TDB}(REF).

4.2 The archival *XMM-Newton* observations of CAL 83

To date, CAL 83 has been observed by *XMM-Newton* 23 times. These observations are summarized in Table 4.2, with the corresponding X-ray instrumental configurations in Table 4.3. The principal investigators (PIs) of these proposals were F. Jansen (0123510101), T. Lanz (0500860201 to 0500860901) and R. Schwarz (0506530201 to 0506531701). Four of the observations were during an X-ray off-state, i.e. the source was not detected above the background level: 0500860701, 0500860801, 0506530301 and 0506530401. This leaves 19 observations of CAL 83 during an X-ray on-state, which were consequently analysed as described in the rest of this chapter. The reductions were performed by following standard data reduction procedures with the *XMM-Newton* SAS Version 13.0.1.

Details pertaining to existing publications making use of these data have already been provided in §3.1, and can be summarized as follows: Spectral and timing analysis of observation 0123510101 have been published by Lanz et al. (2005) and Schmidtke and Cowley (2006) respectively. Rajoelimanana et al. (2013) considered all the other observations, and correlated the long-term X-ray variability with the optical variability as obtained from MACHO and OGLE light curves. These authors also performed blackbody fits to the X-ray data. No detailed WD model atmosphere fitting has yet been applied to the more recent RGS data (2007-2009). To date, the OM data has not been published elsewhere.

In particular, no other authors have yet published the results of detailed timing analysis of these X-ray lightcurves on time-scales of <100 s, which could potentially reveal the spin period of a highly spun-up WD, or possibly rapid non-radial WD pulsations. Schmidtke and Cowley (2006) used a binning of 200 s for their lightcurves of observation 0123510101, and if they also investigated shorter periods it was not mentioned. Despite the discovery in *Chandra* data of the 38.4 min period by these authors, no search for similar periods in the more recent (2007-2009) data has been reported.

¹<http://xmm.esac.esa.int/sas/>

Table 4.2: Archival *XMM-Newton* observations of CAL 83. The EPIC and RGS exposure specifications are listed in Table 4.3. The OM exposures were obtained in Imaging Mode, and are listed in the format Filter(Number of exposures).

Observation ID	Start date & time (UT)	Duration (s)	EPIC, RGS exposures	OM exposures	X-ray state
0123510101	2000-04-23 07:34:01	45021	A, E	None	Bright
0500860201	2007-05-13 22:03:32	11949	B, E	U(4)	Bright
0500860301	2007-07-06 23:31:52	10920	B, E	U(3)	Bright
0500860401	2007-08-21 15:11:21	7915	B, E	U(2)	Bright
0500860501	2007-10-05 23:49:21	17615	B, E	U(3)	Bright
0500860601	2007-11-24 21:10:14	23173	B, E	U(5)	Bright
0500860701	2008-01-16 13:24:33	10914	B, E	U(2)	Off/Low
0500860801	2008-03-10 10:14:38	6916	B, E	None	Off/Low
0506530201	2008-03-20 00:33:22	7715	C, E	B(2),UVW1(1), UVW2(2)	Faint
0506530301	2008-04-03 18:40:55	14918	C, E	B(3),UVW1(2), UVW2(4)	Off/Low
0506530401	2008-04-11 06:02:24	5914	D, E	B(2),UVW1(1), UVW2(2)	Off/Low
0506530501	2008-04-16 12:32:01	10873	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506530601	2008-04-17 13:40:13	11217	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506530801	2008-04-19 06:43:35	5916	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506530901	2008-04-20 22:38:43	11617	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0500860901	2008-04-21 02:10:24	12716	B, E	U(4)	Faint
0506531001	2008-04-21 18:47:57	9076	D, E	B(1),UVW1(1), UVM2(1)	Faint
0506531201	2008-04-23 11:20:22	7418	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506531301	2008-04-25 08:13:32	9614	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506531401	2008-04-29 00:40:23	14117	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506531501	2008-08-12 14:50:27	6918	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Bright
0506531601	2008-09-17 11:10:21	6817	D, E	B(1),UVW1(1), UVM2(1),UVW2(1)	Faint
0506531701	2009-05-30 08:00:48	46115	D, E	UVW1(1),UVM2(6), UVW2(3)	Bright

Table 4.3: Observing sequences applicable to Table 4.2. Data modes: FF=Full Frame, EFF=Extended FF, SW=Small Window and SES=Spectroscopy. EPN refers to EPIC pn. Number of exposures and optical blocking filters used are indicated in parentheses.

Instrument	Configuration	Exposures
EPIC	A	EPN EFF(3:thin,medium,thick), MOS1 FF(3:thin,medium,thick), MOS2 FF(3:thin,medium,thick)
EPIC	B	EPN SW(1:medium), MOS1 SW(1:medium), MOS2 SW(1:medium)
EPIC	C	EPN SW(1:thin), MOS1 SW(1:medium), MOS2 SW(1:medium)
EPIC	D	EPN SW(1:thin), MOS1 SW(1:thin), MOS2 SW(1:thin)
RGS	E	RGS1 SES(1), RGS2 SES(1)

4.3 The X-ray and optical lightcurves

4.3.1 The EPIC lightcurves

The necessary calibrations were applied to the EPIC pn and EPIC MOS event files by using the SAS tasks **epproc** and **emproc** respectively. Barycentric corrections were applied to the time stamps by using the **barycen** task. For each observation, a circular aperture centred on the target was defined as the target region. An aperture radius of 40 arcsec was used, which is estimated to include $\gtrsim 90\%$ of the target photons for the EPIC detectors (*XMM-Newton* Users Handbook, 2013, pp. 12-13).

For all the EPIC pn data, a rectangular background region with an area 2.8 times larger than the target aperture was defined in a source-free region on the same CCD as the target. As the field of view of the MOS detectors in Small Window mode is very small, the available off-target regions on the same CCD are limited, and a circular source-free background region with a radius of 15 arcsec was used, with an area a fraction of 0.14 of the source region area. There were a few exceptions: the MOS data of observation 0123510101 was recorded in Full Frame mode, and here a background annulus outside the target aperture was defined, with an area 2.8 times that of the target aperture. For observations 0506530801 and 0506531601, the soft X-ray background counts in the MOS detectors were so sparse that enough background counts could not be found in a 15 arcsec radius off-source region on the same CCD: for these observations, a rectangular region with the same size as the pn background region was defined on an adjacent MOS CCD.

Using only counts falling within the defined regions, the SAS task **evselect** was utilized to extract lightcurves from the EPIC event files. For EPIC MOS data, event patterns ≤ 12 were used, while for EPIC pn data, only event patterns ≤ 4 were selected² (SAS Users Guide, 2014, pp. 46-48). The lightcurve extracted from the target aperture still included background events as well, and is therefore actually a “source+background” lightcurve. The source+background and the background lightcurve (the latter being extracted from the defined background region) were provided as input to the task **epiclccorr**, which created a final, background subtracted source lightcurve, and also performed additional corrections for effects like vignetting, bad pixels, chip gaps, the shape of the point spread function (PSF), filter transmission, quantum efficiency, dead time and good time intervals. Lightcurves were extracted in this manner for photon events in three different energy ranges:

²A pattern parameter ranging from 0 upwards (up to 12 for EPIC pn and up to 32 for MOS) is assigned to each detected event by an on-board recognition scheme, and describes the shape of a detected event on the CCD. Pattern 0 represents a single pixel event, with counts above a defined threshold, while the other patterns involve various different distribution shapes around the central pixel of other pixels also with counts above the threshold.

broadband (0.15-1.0 keV), soft (0.15-0.25 keV) and hard (0.25-1.0 keV) lightcurves. A variety of bin sizes ranging from 1 to 300 s were used.

During observation 0123510101, three exposures were obtained with each EPIC detector, with thin, medium and thick optical blocking filters respectively. The gaps between the exposures were almost 5 ks, therefore they were analysed separately, and were assigned observation IDs of 0123510101A, 0123510101B and 0123510101C. As there are 19 on-state observations, this subdivision of observation 0123510101 yields a total of 21 on-state datasets to be considered. For the other EPIC datasets, either a thin or a medium filter was used (see Tables 4.2 and 4.3). It must be kept in mind that the choice of optical blocking filter also influences the X-ray count rate somewhat, especially at soft energies.

To estimate the effect of the optical blocking filter, the *Chandra* proposal planning tool PIMMS v4.7b was used, which provides general functionality for conversions between source flux and count rate for a variety of X-ray instruments³. Firstly, an estimate of the source parameters was obtained by using Sherpa⁴ (a spectral analysis package forming part of the CIAO 4.6 software) to fit an absorbed blackbody model to the *Chandra* HRC-S/LETG on-state observation of CAL 83 (obs ID 1900). The latter observation is more suited than the *XMM-Newton* data for obtaining spectral parameters for the source, because it extends to lower energies.

When fixing the column density to $6.5 \times 10^{20} \text{ cm}^{-2}$ (Gänsicke et al., 1998), the best fit effective temperature was 45 eV, and the absorbed flux in the 0.07-10 keV band was $F_{\text{abs}} \sim 7.5 \times 10^{-12} \text{ erg cm}^{-2} \text{ s}^{-1}$. This yielded an unabsorbed bolometric luminosity of $L_{\text{bol}} \sim 5.6 \times 10^{37} \text{ erg s}^{-1}$. This temperature is comparable to the $47.4 \pm 2.2 \text{ eV}$ obtained by Lanz et al. (2005) from their WD model atmosphere analysis. Although a blackbody is a very crude estimation for an SSS spectrum, and may lead to an overestimation of the *bolometric* flux, the best fit blackbody provides a flux estimate over the narrow supersoft X-ray band that is adequate for the current approximation.

These blackbody parameters were then adopted as “standard” on-state parameters of CAL 83, and were provided as input to PIMMS. The predicted EPIC pn count rates with thin, medium and thick optical blocking filters were 8.4, 6.8 and 3.2 counts s^{-1} respectively in the 0.15-1.0 keV band. Thus, for pn data, multiplicative factors of $8.4 \div 6.8 = 1.2$ and $8.4 \div 3.2 = 2.6$ could be used to convert count rates obtained with the medium and thick filters respectively to an equivalent count rate with the thin filter. For EPIC MOS, the estimations were 1.2, 0.99 and 0.54 counts s^{-1} , yielding similar factors of 1.2 and 2.2. These factors are of course

³<http://cxc.harvard.edu/toolkit/pimms.jsp>

⁴<http://cxc.harvard.edu/sherpa4.4/>

model-dependent. Keeping in mind that the long-term X-ray flux changes in CAL 83 are probably related to temperature modulations, these factors may not be valid for a wide range of count rates.

Alternatively, one can assume for the moment that the average X-ray luminosity of CAL 83 was identical for each of the three EPIC segments in observation 0123510101. The mean count rates for the three segments were 6.5, 5.5 and 2.8 counts s⁻¹ for pn, and 2.6, 2.3 and 0.65 counts s⁻¹ for MOS. These yield multiplicative factors of 1.2 and 2.3 for pn, and 1.3 and 2.4 for MOS, for conversion from medium and thick filter counts respectively to “thin filter counts”. These values compare quite well with the factors predicted by PIMMS. Subsequently, the factors derived from observation 0123510101 were applied to all the EPIC observations obtained with medium and thick optical blocking filters, to yield the approximate count rate that would have been obtained if a thin filter was used. The benefit of this transformation is that the EPIC count rates for different observations can now be compared directly. Because it is a simple linear transformation of the count rates and errors, it will have no influence on the period analysis presented later in this chapter.

The MOS1 and MOS2 exposures were obtained over almost identical time intervals. Upon extraction with **evselect**, identical time bins were therefore defined for the MOS1 and MOS2 lightcurves, trimming off the occasional few seconds at the edges of the lightcurves where they did not coincide. After processing with **epiclccorr**, the MOS1 and MOS2 lightcurves were averaged to obtain a single MOS lightcurve for each observation. A weighted mean was used, with the weight equal to the inverse of the mean error in the count rate of each respective lightcurve. The MOS exposure intervals were usually displaced with respect to the pn exposure intervals by a more significant amount, ranging from a few seconds up to about an hour, which is the reason why the MOS lightcurves were not combined with the pn lightcurves.

Consequently, a hardness ratio lightcurve was also calculated from each set of soft and hard lightcurves. The hardness ratio was defined as

$$\text{HR} = \frac{H - S}{H + S} , \quad (4.6)$$

where S and H represent the count rates in the soft and hard bands respectively.

Because the EPIC pn detector has a significantly larger effective area than even the combination of the two EPIC MOS detectors (see Fig. 4.2), the signal-to-noise ratio (S/N) of the pn lightcurves was significantly larger. Although all the analyses reported here were performed for both the pn and MOS lightcurves, the approach in this chapter will mostly be to present the pn results, with occasional supplemen-

Table 4.4: A summary of the amount of variability in the EPIC pn lightcurves of CAL 83. The count rates have been corrected to “thin filter counts” (see text).

Obs ID	Average CR (counts s ⁻¹)	Standard deviation σ_{CR} in CR (counts s ⁻¹)	σ_{CR} as % of mean CR	Maximum modulation %	X-ray state
0123510101A	6.470 $\pm$ 0.029	0.321	5	12	Bright
0123510101B	6.472 $\pm$ 0.032	0.724	11	29	Bright
0123510101C	6.472 $\pm$ 0.045	0.783	12	24	Bright
0500860201	7.324 $\pm$ 0.036	0.761	10	25	Bright
0500860301	6.764 $\pm$ 0.036	0.588	9	27	Bright
0500860401	5.208 $\pm$ 0.037	0.588	11	28	Bright
0500860501	5.034 $\pm$ 0.028	1.061	21	55	Bright
0500860601	6.445 $\pm$ 0.026	1.013	16	39	Bright
0500860901	0.234 $\pm$ 0.008	0.097	41	133	Faint
0506530201	0.619 $\pm$ 0.013	0.160	26	80	Faint
0506530501	1.680 $\pm$ 0.025	0.248	15	32	Faint
0506530601	1.426 $\pm$ 0.015	0.436	31	85	Faint
0506530801	0.350 $\pm$ 0.012	0.133	38	94	Faint
0506530901	0.403 $\pm$ 0.008	0.171	42	114	Faint
0506531001	0.507 $\pm$ 0.012	0.153	30	76	Faint
0506531201	0.166 $\pm$ 0.008	0.070	42	95	Faint
0506531301	0.438 $\pm$ 0.010	0.169	39	124	Faint
0506531401	0.249 $\pm$ 0.006	0.170	68	216	Faint
0506531501	7.846 $\pm$ 0.045	0.696	9	25	Bright
0506531601	0.101 $\pm$ 0.007	0.106	105	347	Faint
0506531701	7.687 $\pm$ 0.017	0.793	10	31	Bright

tation by the similar, independent results obtained from the MOS lightcurves. Due to the much lower throughput, the RGS detectors are not ideal for timing, and were therefore not included in the current analysis. However, follow-up studies should definitely include the fitting of line-blanketed WD atmosphere models to the high-resolution RGS spectra.

The EPIC pn broadband lightcurves are presented in Fig. 4.5, together with the HR lightcurves, as well as the photometric measurements obtained from the Optical Monitor (see the next section for a discussion of the OM data). It is evident that there is significant variability in the X-ray count rate and hardness ratio on time-scales of minutes to hours. The MOS lightcurves exhibit similar modulations, albeit with a higher noise level. The variability encompasses a wide range of time-scales, amplitudes and modulation shapes. One can also see that the X-ray count rate and hardness ratio follow the same trends in many of the observations. This will be explored in more detail in §4.3.3.

In order to quantify the variability in the EPIC pn lightcurves, a few simple calculations were performed. The standard deviation of the counts in each lightcurve was calculated, and expressed as a percentage of the mean counts. The data point in each lightcurve with the largest deviation from the mean was also identified, and the difference converted to a percentage of the mean to obtain a “maximum modulation percentage” value. The results of these calculations are listed in Table 4.4. Clearly, the amount of variability in the lightcurves vary greatly, with standard deviations between 5 and 105% of the mean count rate.

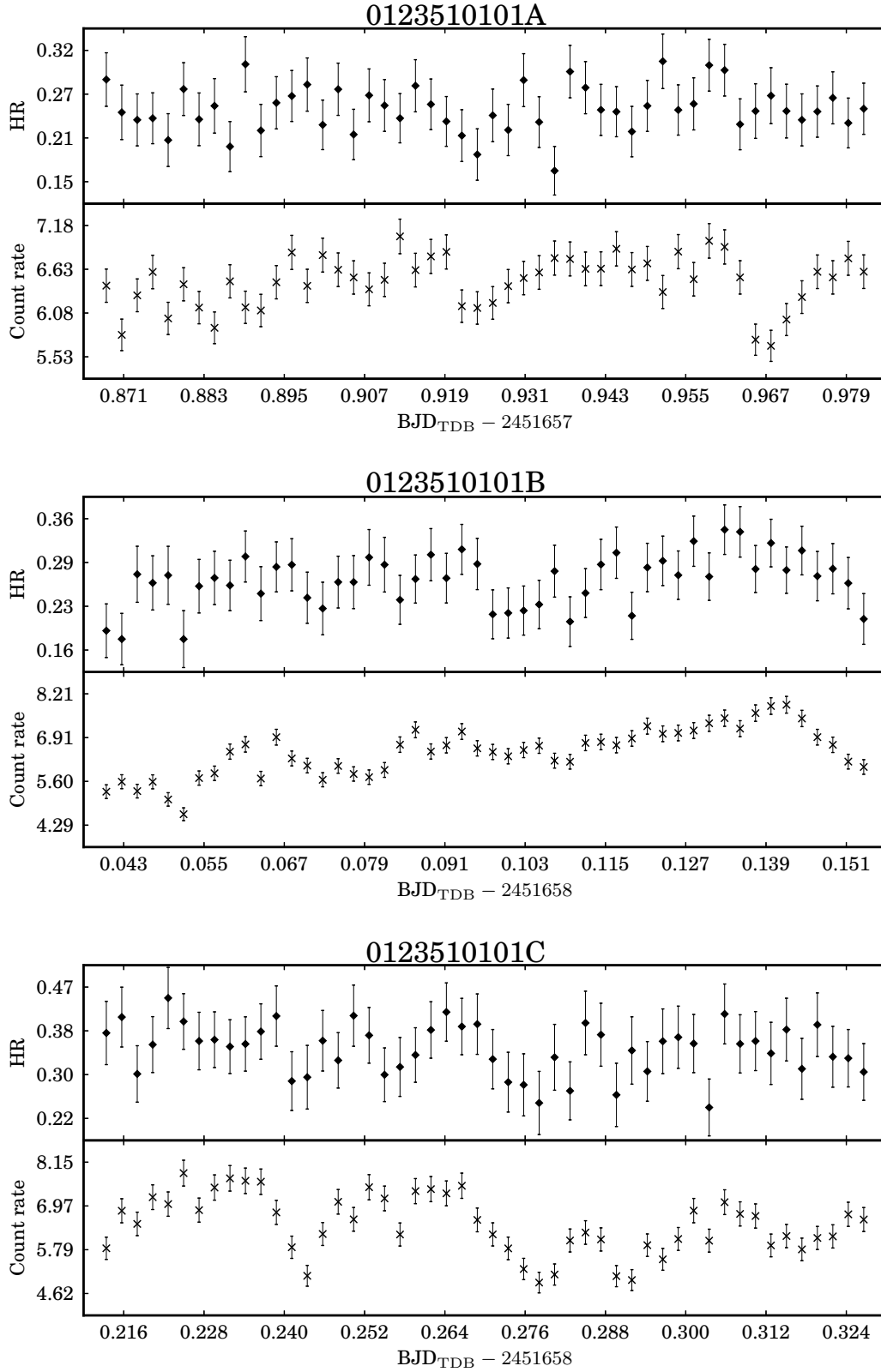


Fig. 4.5: The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. The optical blocking filters that were used were thin (A), medium (B) and thick (C) for the three respective segments.

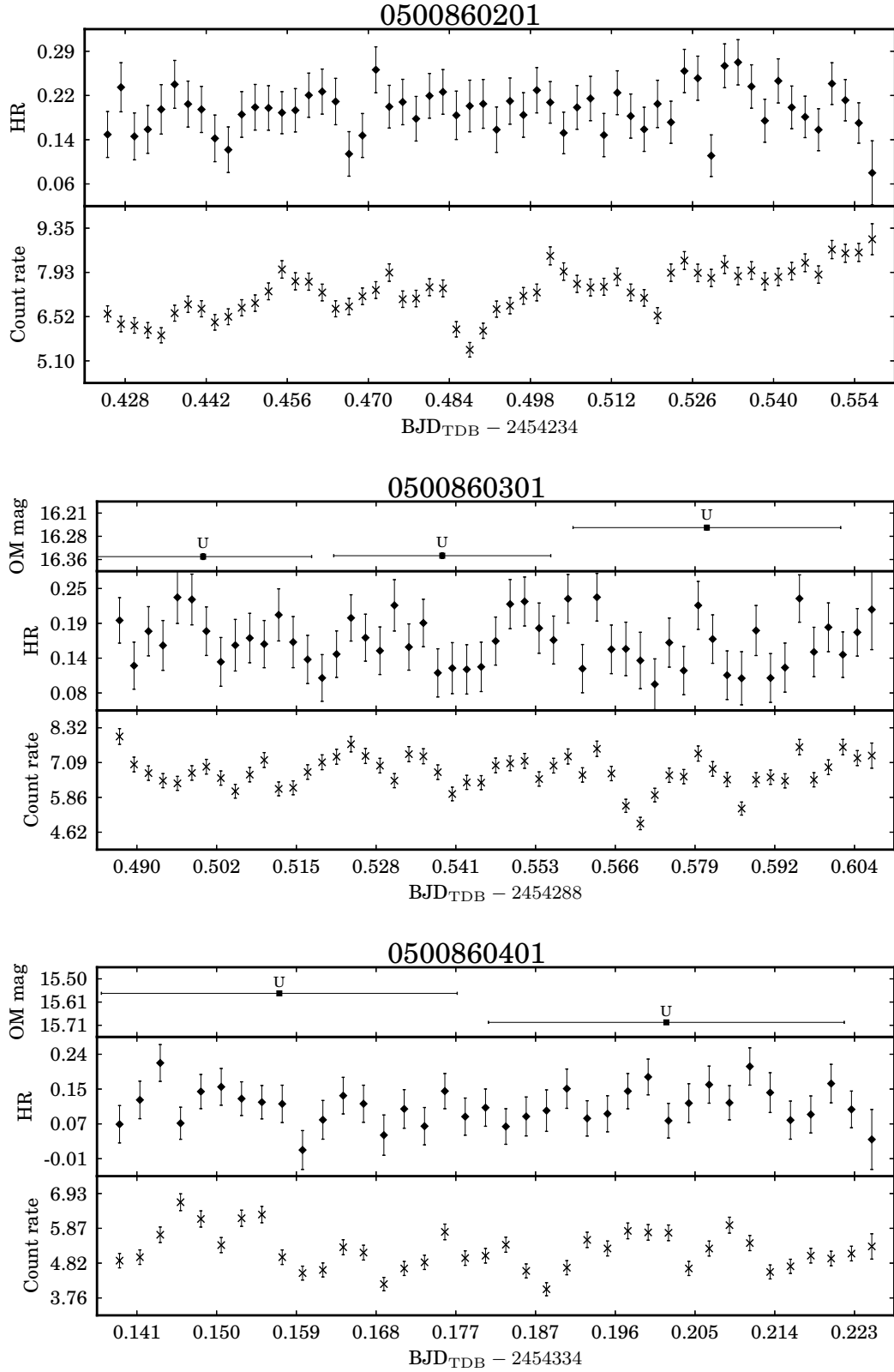


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

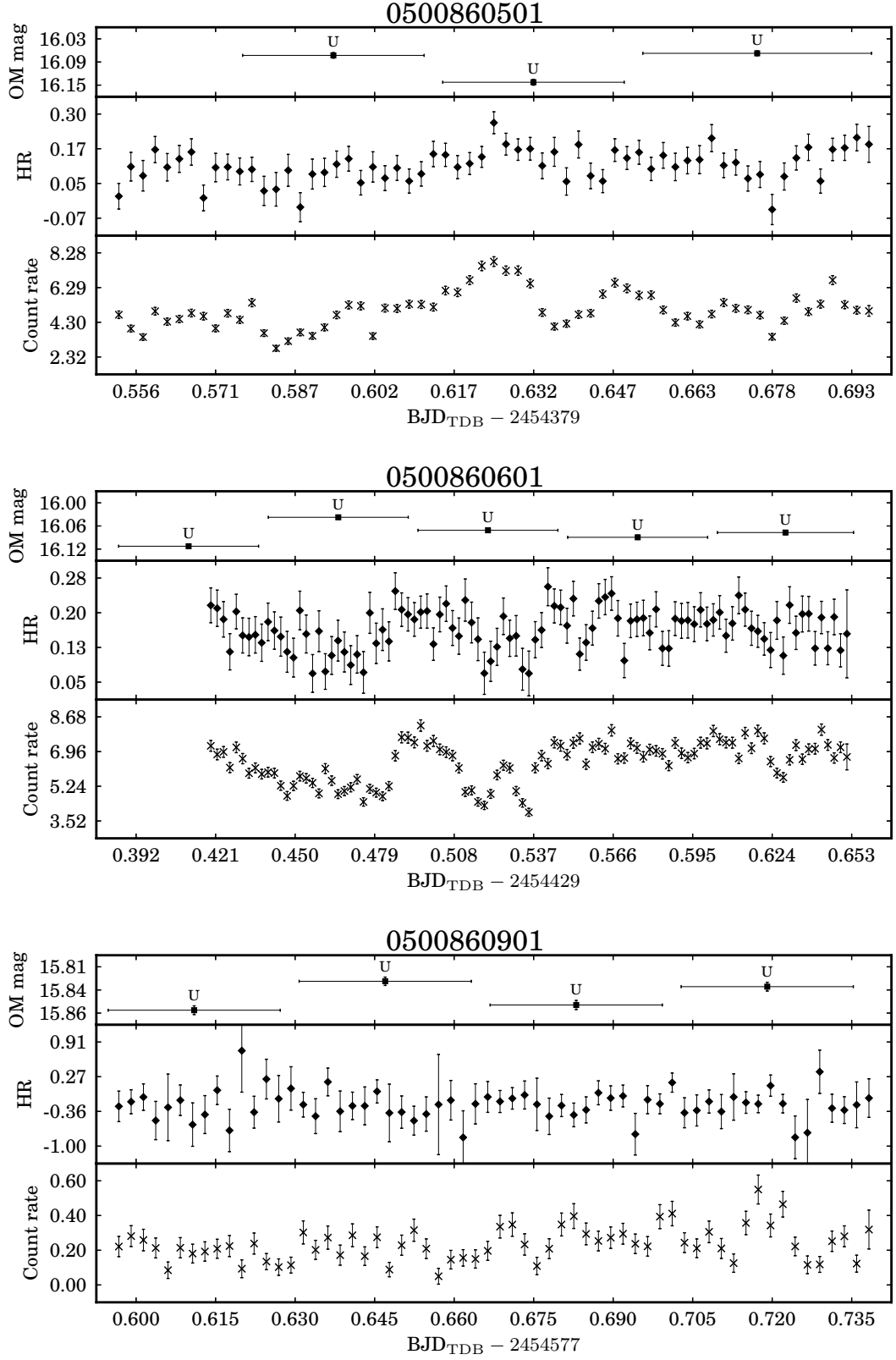


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

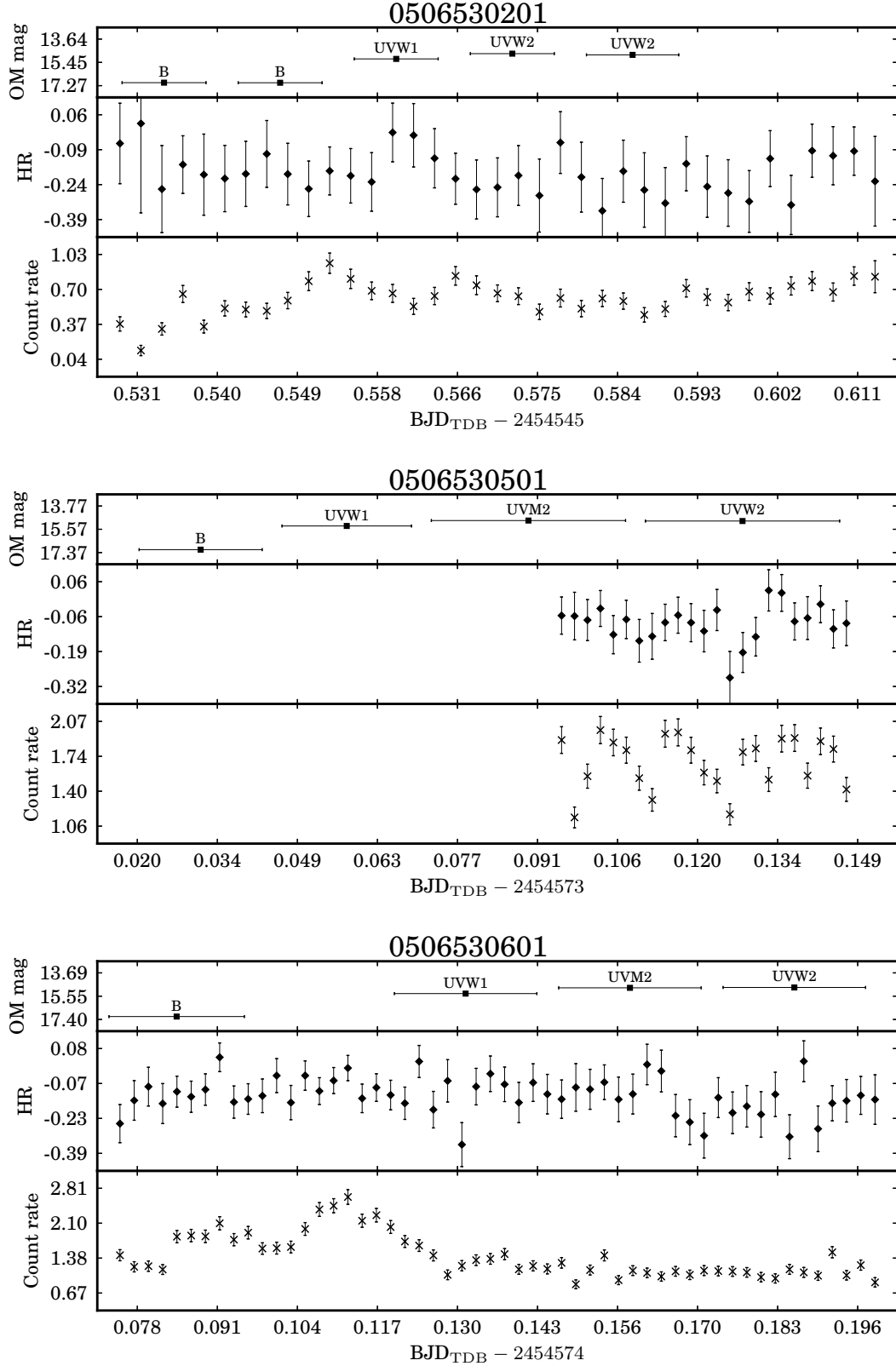


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

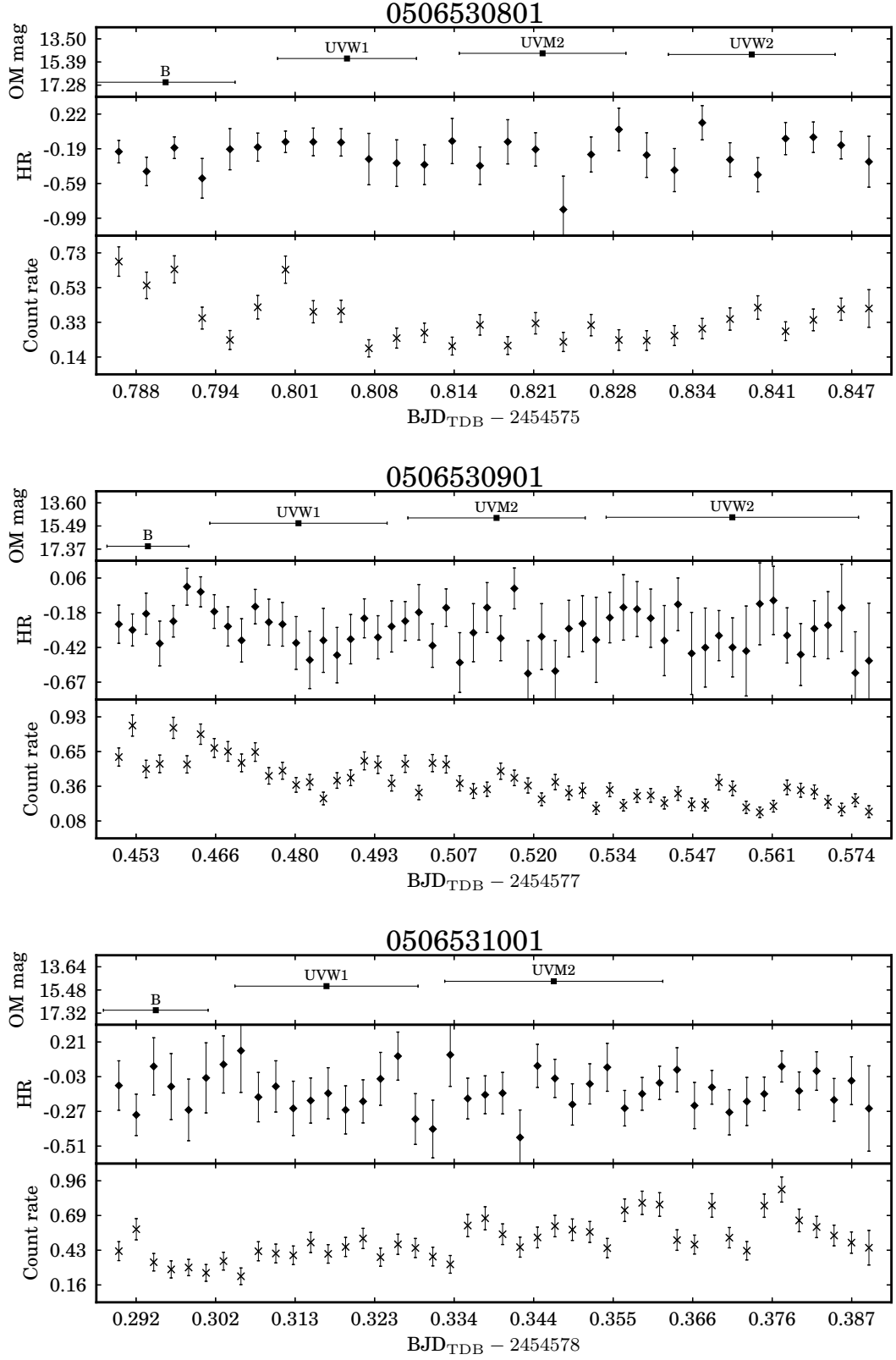


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

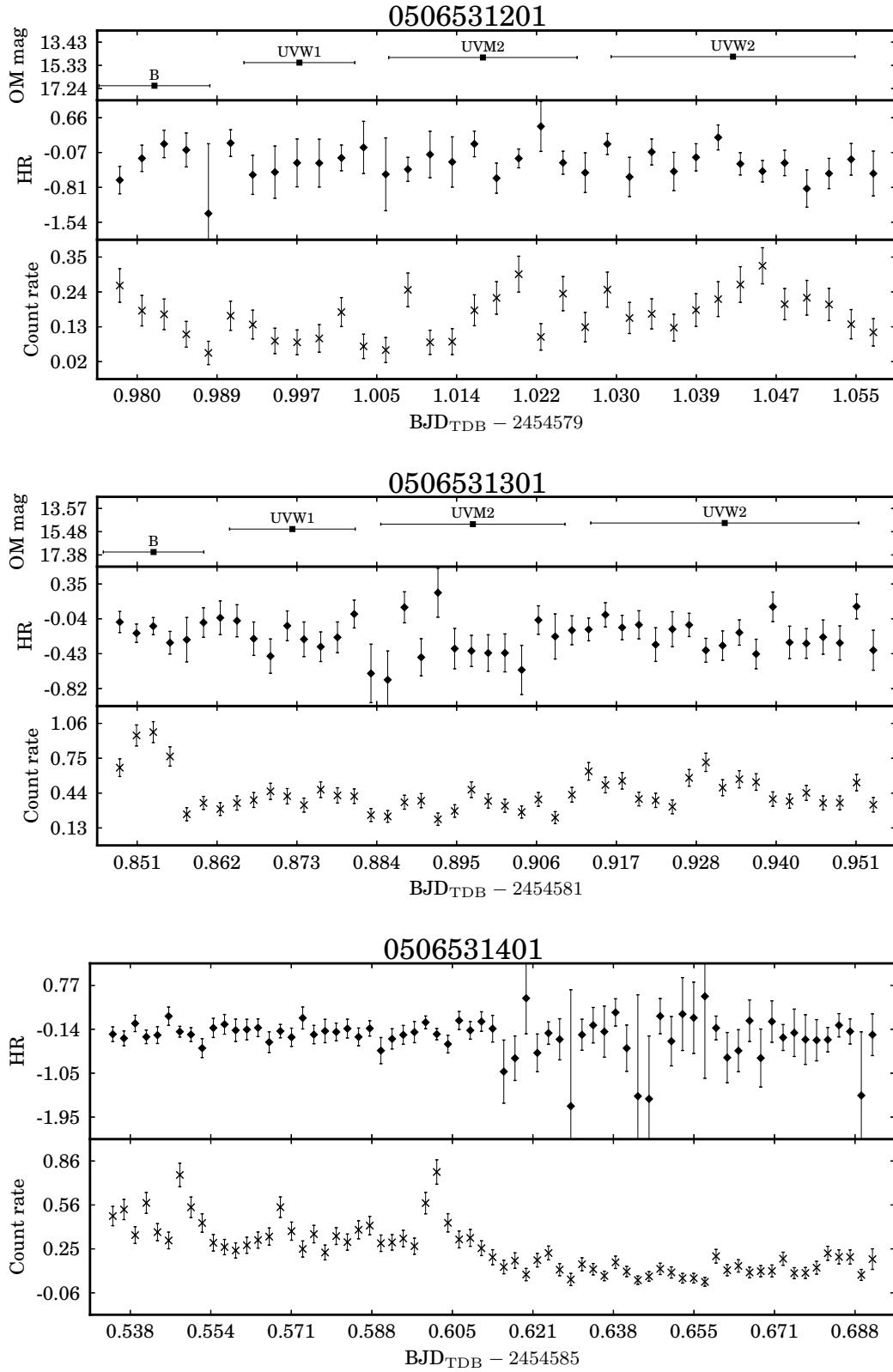


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

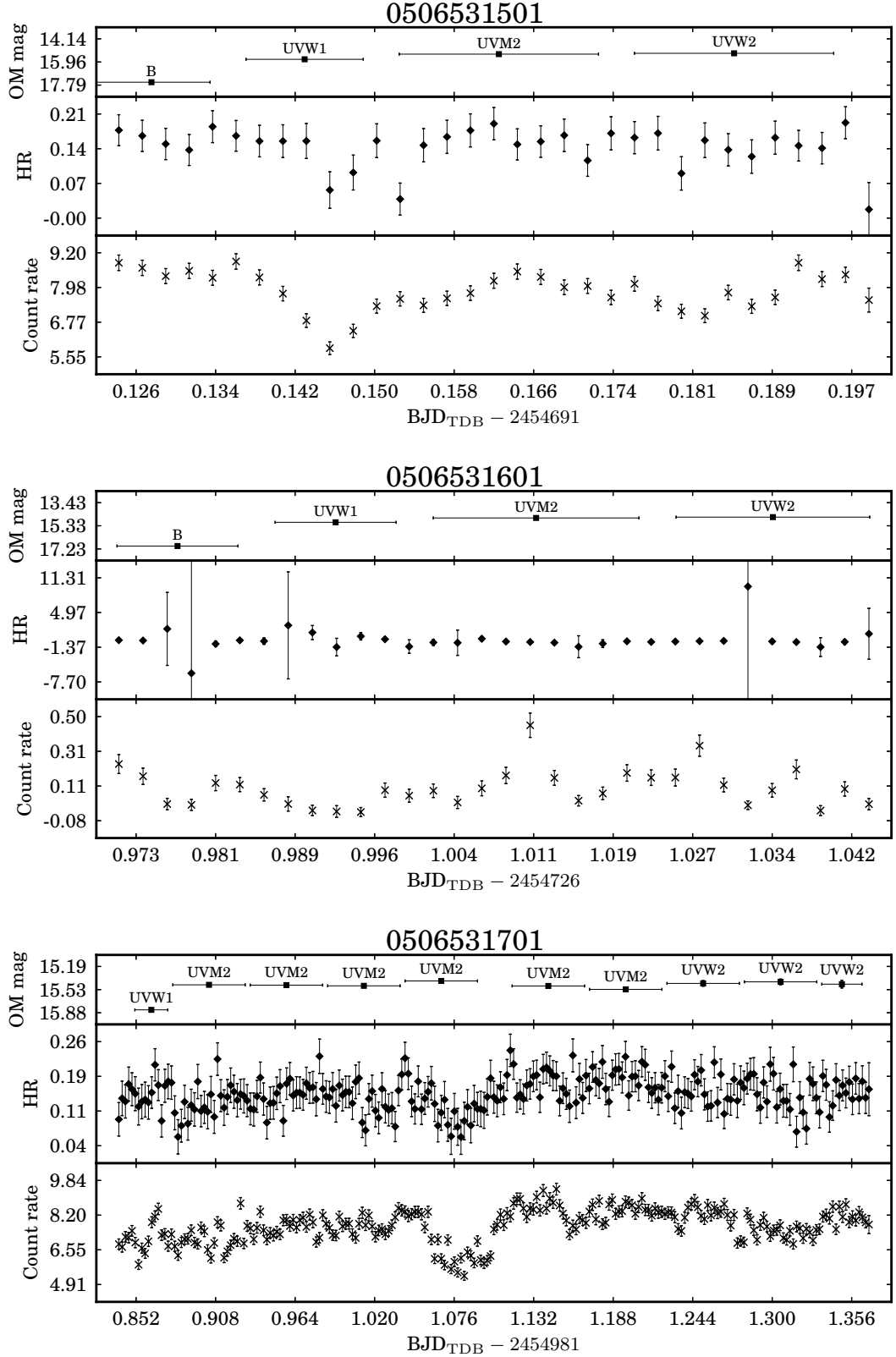


Fig. 4.5: (continued) The EPIC pn broadband lightcurves of the archival *XMM-Newton* observations of CAL 83, together with the corresponding EPIC pn hardness ratios (HR). See Eq. (4.6) for the definition of HR. Each X-ray data point represents a time bin of 200 s. Where available, the OM instrumental magnitudes are also plotted, with the horizontal error bars indicating the OM exposure intervals (see Table B.1).

4.3.2 Photometry with the Optical Monitor

As indicated in Table 4.2, all of the observations except 0123510101 and 0500860801 also include OM data. The processing of the OM data files were performed with the SAS task **omichain**. The instrumental magnitude of CAL 83 from each of the OM images was determined with SAS, and these are given in Table B.1 in Appendix B. The OM instrumental magnitudes are also presented in graphical form in Fig. 4.5, together with the EPIC pn count rate (CR) and hardness ratio (HR) lightcurves. Unfortunately, the OM data files of observations 0500860201 and 0506531401 could not be processed by SAS, and appear to be corrupted.

In observations 0500860301, 0500860401, 0500860501, 0500860601, 0500860901 and 0506530201, more than one exposure with the same filter was obtained during a single pointing, and the magnitude variation can be investigated. To enable an easier comparison with the X-ray lightcurves, the EPIC pn CR and HR were averaged over the time intervals of the OM exposures. These values, together with the OM magnitudes, are shown in Fig. 4.6.

It is evident that the U- and UVM2-filter magnitudes exhibited very small but statistically significant modulations during single observations. Although the CR and HR values confirm the correlation inferred from Fig. 4.5, no consistent correlation or anti-correlation of these quantities with the optical/UV magnitudes was found within the individual observations. The B- and UVW2-filter magnitudes of observations 0506530201 and 0506531701 did not vary significantly.

In addition to the individual instrumental magnitude measurement for each exposure, a combined instrumental magnitude for each set of images taken with the same filter was obtained with SAS, as well as the specific flux and the AB magnitude. These measurements are given in Table B.2 in Appendix B. The BJD_{TDB} in this table was obtained by simply averaging the mid-exposure times from Table B.1 for the relevant combined photometric measurements.

In the AB system, the monochromatic magnitude is simply (e.g. Tonry et al., 2012)

$$m_{\text{AB}}(\nu) = -2.5 \log \left(\frac{f_{\nu}}{3631 \text{ Jy}} \right), \quad (4.7)$$

with f_{ν} the specific flux from the target at frequency ν in units of $\text{erg s}^{-1} \text{ cm}^{-2} \text{ Hz}^{-1}$, and 1 Jansky (Jy) = $10^{-23} \text{ erg s}^{-1} \text{ cm}^{-2} \text{ Hz}^{-1}$. The value of the constant in the equations above sets the AB magnitude of Vega at 548 nm to 0.03, which is also the V magnitude of this source in the Vega system. However, practical measurements are always over a certain wavelength range (bandpass), and the “bandpass

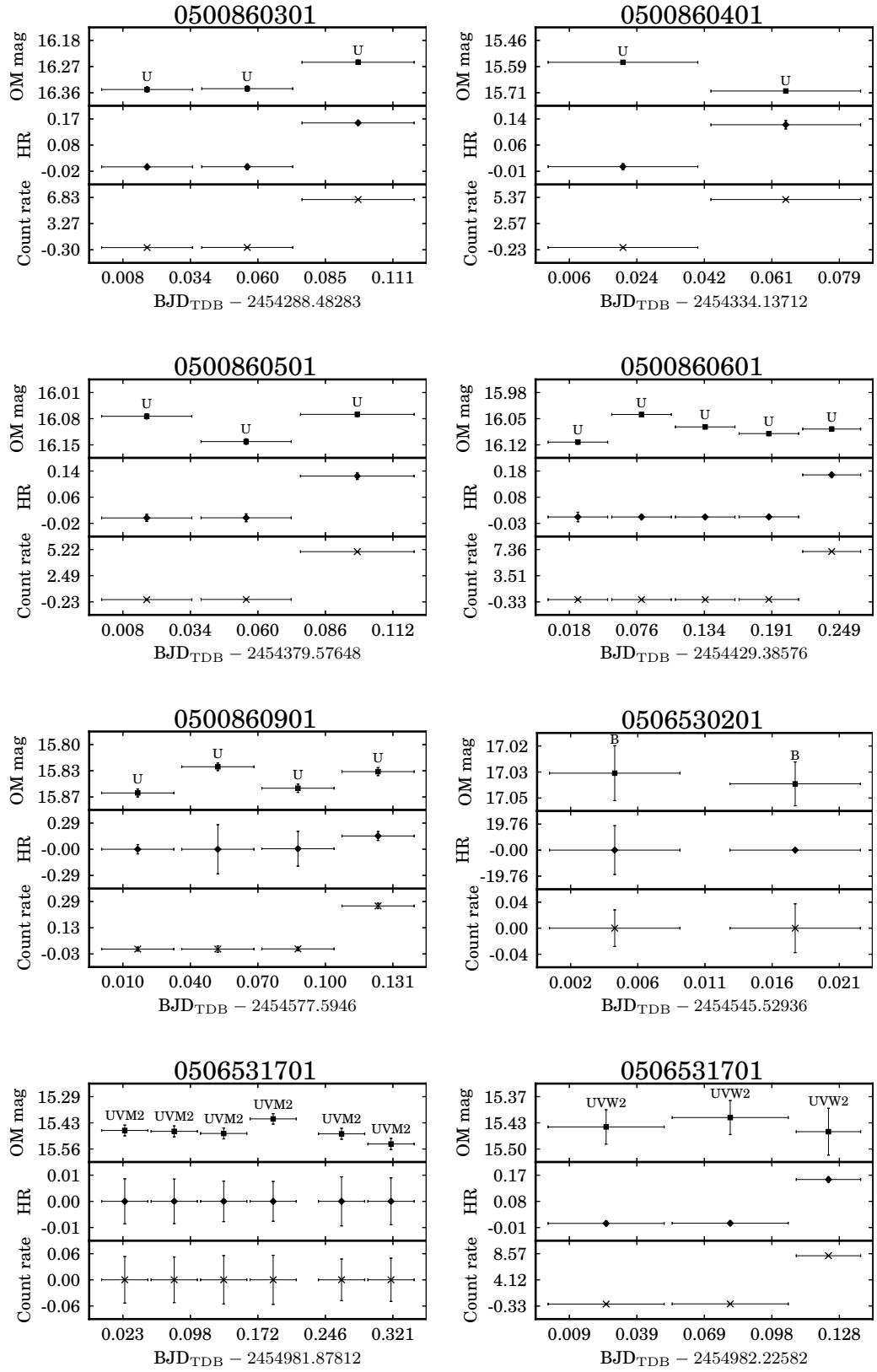


Fig. 4.6: The OM instrumental magnitudes for observations containing more than one exposure with the same OM filter, plotted with the pn count rate and hardness ratio HR averaged over the length of each OM exposure interval. Vertical error bars were included, although they are too small to discriminate from the data points in some cases.

Table 4.5: Effective wavelengths of OM filters.

Filter	Effective wavelength λ_{eff} (Å)
B	4506
U	3441
UVW1	2908
UVM2	2311
UVW2	2119

AB magnitude” is defined in a similar manner:

$$m_{\text{AB}} = -2.5 \log \left(\frac{\int f_{\nu} (h\nu)^{-1} A(\nu) d\nu}{\int 3631 \text{ Jy} (h\nu)^{-1} A(\nu) d\nu} \right) , \quad (4.8)$$

with $A(\nu)$ the “net capture cross-section” of the detector to photons with frequency ν , which depends on the system properties, with the filter throughput being a very important factor. Because

$$f_{\nu} = \frac{\lambda^2}{c} f_{\lambda} , \quad (4.9)$$

where $\lambda = c/\nu$ is the wavelength related to ν , and f_{λ} the specific flux in $\text{erg s}^{-1} \text{ cm}^{-2} \text{ Å}^{-1}$, Eq. (4.7) can be rewritten in terms of wavelength as

$$m_{\text{AB}}(\lambda) = -2.5 \log \left(\frac{\frac{\lambda^2}{c} f_{\lambda}}{3631 \text{ Jy}} \right) , \quad (4.10)$$

allowing Eq. (4.8) to be rewritten in a similar way. It is often useful to consider a photometric measurement as a specific flux measurement f_{λ} (or f_{ν}) at a specific wavelength (or frequency). However, the bandpass of a filter encompasses a whole range of wavelengths, so which wavelength should be used? The “effective wavelength” λ_{eff} is introduced, so that after calculating m_{AB} with Eq. (4.8), λ_{eff} is the appropriate wavelength to be used in Eq. (4.10) to relate m_{AB} to f_{λ} . An effective frequency $\nu_{\text{eff}} = \frac{c}{\lambda_{\text{eff}}}$ may also be determined. The values of the effective wavelengths of the OM filters used in the CAL 83 observations are listed in Table 4.5. These were calculated by comparing the magnitude and flux measurements obtained with **omichain**.

All the AB magnitudes measured from the OM data will be presented in graphical form in §4.6 together with the OGLE-III lightcurve. The corresponding specific flux values from observations containing measurements with more than one filter are plotted in Fig. 4.7 versus the effective wavelengths given in Table 4.5, to illustrate the spectral energy distribution of the source in the near-UV waveband. The blue nature of the spectrum is immediately evident, with the tail in the UV presumably extending from a peak at even shorter wavelengths. The long-term variability in the near-UV flux will be illustrated more clearly in §4.6.

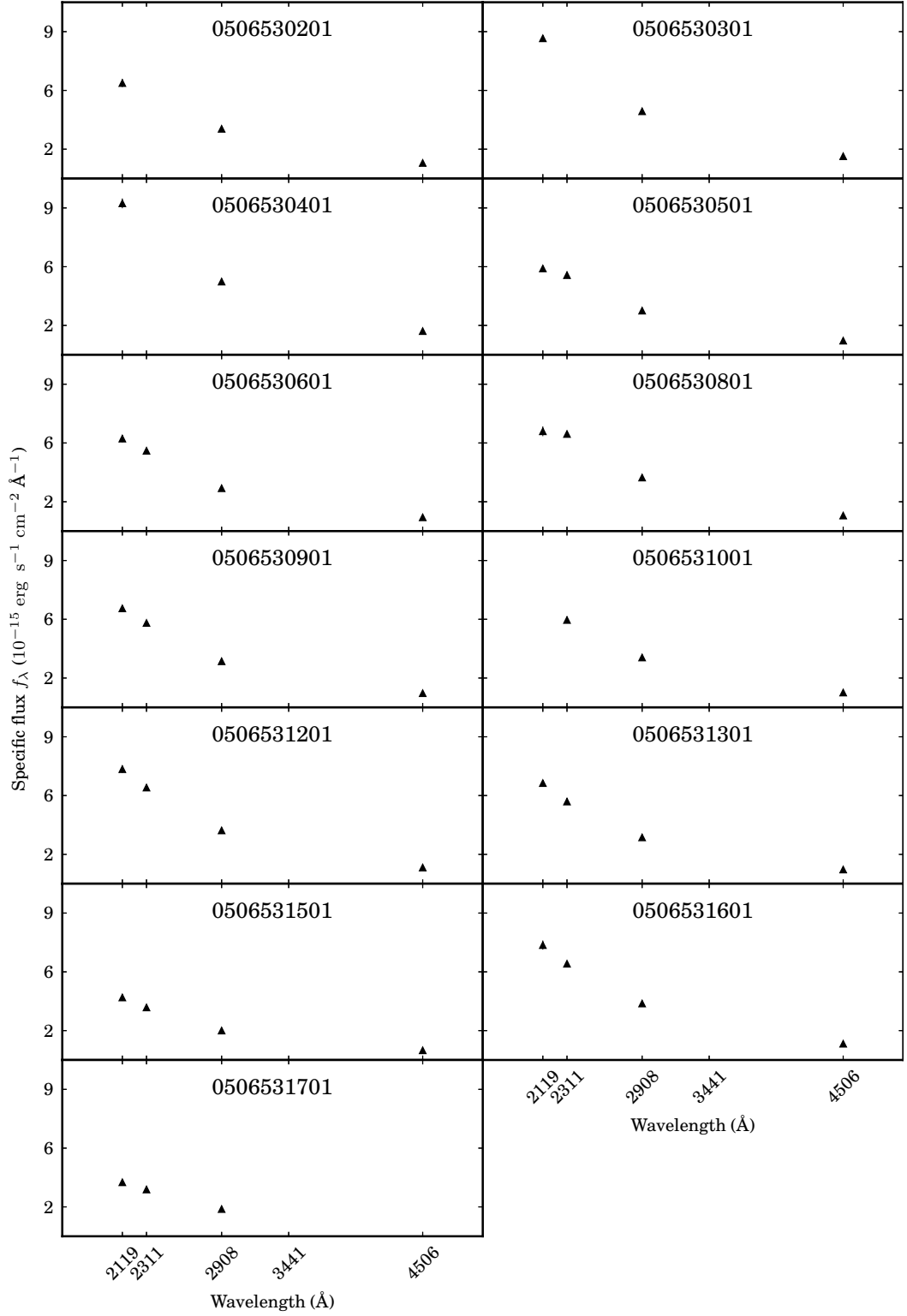


Fig. 4.7: The spectral energy distribution of CAL 83 in the near-UV, as derived from the *XMM-Newton* observations containing OM data taken with more than one filter. Error bars are included, although they are smaller than the data markers in most cases.

4.3.3 Correlation of the X-ray count rate and hardness ratio

As mentioned in the previous section, inspection of the lightcurves in Fig. 4.5 reveals an unmistakable correlation between the count rate (CR) and hardness ratio (HR) in many of the observations. In order to quantify this apparent correlation, a simple *rank correlation* (also known as a *nonparametric* correlation) was performed, based on the method described by Press et al. (2007, pp. 748-751).

Consider two quantities a and b that are measured simultaneously, yielding a lightcurve for each of them. The lightcurves consist of measurements a_i and b_i respectively at identical times t_i , where $i = 1, 2, 3, \dots, N$, in other words N represents the number of data points in each lightcurve. The first step is to “rank” each set of measurements a_i and b_i . First, the a_i ’s are arranged from small to large, and the exact same rearrangement is applied to the b_i values, regardless of the actual values of the b_i ’s. The correlation information is therefore not lost, because the n^{th} point in the rearranged a_i lightcurve will still represent a measurement obtained simultaneously with the n^{th} point in the rearranged b_i lightcurve.

Then, each a_i is replaced by its rank A_i among all the other a_i ’s in the sample. If all the different a_i ’s are unique, the resulting ranks will be the integers $1, 2, 3, \dots, N$. If some of the a_i ’s have the same values, the number assigned to these “ties” is the mean of the ranks they would have had if they had been slightly different. In either case, the sum of all the assigned ranks A_i will be equal to the sum of all the integers from 1 to N .

After the ranking of the a_i ’s is completed, the b_i ’s are now sorted from small to large, and the same new rearrangement pattern applied to the A_i ’s. Each b_i value is then replaced by its corresponding rank B_i , as described above. After ranking a_i and b_i , the resulting lists of numbers A_i and B_i are drawn from a perfectly known distribution function, consisting of the integers from 1 to N . This enables one to interpret the statistical significance of the correlation coefficient that will be determined, and is also the reason why the ranks of the values are used rather than the values themselves.

Consequently, the Spearman linear rank-order correlation coefficient was calculated from the ranked lightcurves. This correlation coefficient is given in terms of the assigned ranks A_i and B_i and their mean values of $\bar{A}$ and $\bar{B}$ by

$$r_s = \frac{\sum_i (A_i - \bar{A}) (B_i - \bar{B})}{[\sum_i (A_i - \bar{A})^2]^{1/2} [\sum_i (B_i - \bar{B})^2]^{1/2}} \quad (4.11)$$

for cases where there are no ties. The value of r_s can vary between -1 and $+1$.

A value of -1 indicates a complete anti-correlation between the lightcurves, i.e. a plot of A_i versus B_i would yield a straight line with a negative slope. A value of $+1$ indicates a complete correlation between the lightcurves, i.e. a plot of A_i versus B_i would yield a straight line with a positive slope. If $r_s = 0$, no correlation exists. When a non-zero r_s -value is obtained, its statistical significance can be evaluated by calculating

$$t = r_s \sqrt{\frac{N-2}{1-r_s^2}} . \quad (4.12)$$

The parameter t is distributed approximately as a Student's distribution with $N-2$ degrees of freedom. The value of r_s is also related to another parameter that is regularly used in nonparametric correlation: the *sum-squared difference of ranks*, which is defined as

$$D = \sum_{i=0}^{N-1} (A_i - B_i)^2 \quad (4.13)$$

When there are no ties in the datasets, the relation between r_s and D is

$$r_s = 1 - \frac{6}{N^3 - N} . \quad (4.14)$$

However, when there are ties, the equation takes a more complicated form. If f_k is the number of ties in the k^{th} group of ties among the values of A_i , and g_k the number of ties in the m^{th} group of ties among the values of B_i , the relation is

$$r_s = \frac{1 - \frac{6}{N^3 - N} [D + \frac{1}{12} \sum_k (f_k^3 - f_k) + \frac{1}{12} \sum_m (g_m^3 - g_m)]}{\left[1 - \frac{\sum_k (f_k^3 - f_k)}{N^3 - N}\right]^{1/2} \left[1 - \frac{\sum_m (g_m^3 - g_m)}{N^3 - N}\right]^{1/2}} . \quad (4.15)$$

In the null hypothesis (uncorrelated datasets), the expectation value of D is

$$\overline{D} = \frac{1}{6} (N^3 - N) - \frac{1}{12} \sum_k (f_k^3 - f_k) - \frac{1}{12} \sum_m (g_m^3 - g_m) , \quad (4.16)$$

and its variance

$$\text{Var}(D) = \frac{(N-1)N^2(N+1)^2}{36} \left[1 - \frac{\sum_k (f_k^3 - f_k)}{N^3 - N}\right] \left[1 - \frac{\sum_m (g_m^3 - g_m)}{N^3 - N}\right] . \quad (4.17)$$

The values of D follow an approximately normal distribution, enabling the calculation of its significance levels with a complementary error function. The significance value of D can thus be compared with the significance value of r_s to confirm the reliability of the latter.

Consequently, the method was applied to quantify the correlation and the signif-

Table 4.6: The Spearman linear rank-order correlation coefficient r_s and the sum-squared difference of ranks D , describing the correlation between the X-ray count rate (CR) and hardness ratio (HR) as determined from the EPIC pn lightcurves with 200 s bins. The significance values of both r_s and D are also provided, as well as the mean CR and HR for each observation. The count rates have been corrected to “thin filter counts” (see text).

Obs ID	Average CR (counts s ⁻¹)	Average HR	r_s	Significance of r_s (%)	D	Significance of D (%)
0123510101A	6.470 ± 0.029	0.246 ± 0.005	0.119	58.93	18348	59.49
0123510101B	6.472 ± 0.032	0.266 ± 0.005	0.555	99.997	9262	99.989
0123510101C	6.472 ± 0.045	0.350 ± 0.008	0.517	99.99	10066	99.97
0500860201	7.324 ± 0.036	0.192 ± 0.005	0.205	87.81	25832	87.90
0500860301	6.764 ± 0.036	0.165 ± 0.006	0.437	99.89	13954	99.84
0500860401	5.208 ± 0.037	0.112 ± 0.007	0.336	96.10	6066	95.92
0500860501	5.034 ± 0.028	0.116 ± 0.006	0.448	99.98	23010	99.96
0500860601	6.445 ± 0.026	0.166 ± 0.004	0.510	99.999995	84082	99.99997
0500860901	0.234 ± 0.008	-0.239 ± 0.045	0.095	53.88	35926	54.34
0506530201	0.619 ± 0.013	-0.189 ± 0.025	-0.021	9.75	8612	9.96
0506530501	1.680 ± 0.025	-0.090 ± 0.015	0.316	85.84	1384	86.20
0506530601	1.426 ± 0.015	-0.127 ± 0.012	0.269	95.10	19172	95.00
0506530801	0.350 ± 0.012	-0.248 ± 0.042	-0.055	21.79	3854	22.39
0506530901	0.403 ± 0.008	-0.301 ± 0.027	0.144	70.88	25060	71.29
0506531001	0.507 ± 0.012	-0.127 ± 0.026	-0.098	47.30	15580	47.93
0506531201	0.166 ± 0.008	-0.293 ± 0.076	-0.005	2.17	7174	2.22
0506531301	0.438 ± 0.010	-0.238 ± 0.028	0.153	69.16	13726	69.69
0506531401	0.249 ± 0.006	-0.288 ± 0.071	0.075	45.76	50656	46.16
0506531501	7.846 ± 0.045	0.144 ± 0.006	0.350	95.43	3888	95.25
0506531601	0.101 ± 0.007	-0.019 ± 2.226	-0.027	11.56	5602	11.84
0506531701	7.687 ± 0.017	0.151 ± 0.002	0.467	≫99.999999	1066534	≫99.999999

icance thereof between the EPIC pn CR and HR for each observation separately, using the lightcurves with a bin size of 200 s. By setting $a = \text{CR}$ and $b = \text{HR}$, a PYTHON script based on the C++ code provided by Press et al. (2007, pp. 750-751) was used to calculate the Spearman linear rank-order correlation coefficient r_s and the sum-squared difference of ranks D , as well as their significance levels in each case. The results are presented in Table 4.6. The mean pn count rate and hardness ratio are also provided for each observation.

A correlation between the count rate and hardness ratio was considered to be statistically significant if $0 < r_s \leq +1$, with a significance level $>99.73\%$. As can be seen in Table 4.6, there were 6 datasets obeying this condition: 0123510101B, 0123510101C, 0500860301, 0500860501, 0500860601 and 0506531701. Relaxing the significance requirement to $>95.45\%$ adds observation 0500860401 to the sample, while an additional 6 observations exhibited a correlation significance above 68.27%: 0500860201, 0506530501, 0506530601, 0506530901, 0506531301 and 0506531501.

Five of the on-state observations yielded a negative value for r_s : 0506530201, 0506530801, 0506531001, 0506531201 and 0506531601. At first glance, this might seem like an indication of an anti-correlation. However, upon closer inspection one can see that: (i) these negative values are very close to zero, (ii) they have very low

significance values; in fact, the lowest of all the observations, ranging from 2.2 to 47.3%, effectively placing these measurements in the noise regime, and (iii) the mean count rates during these observations were very low. These “anti-correlations” are therefore not significant.

Finally, the *mean* CRs and HRs for all the observations as listed in Table 4.6 were correlated to investigate the relation between these quantities from several days to several years. The correlation coefficient is $r_s = +0.817$, with a significance of 99.99994%. It is concluded that, in general, the hardness ratio is correlated with the count rate.

After having illustrated the presence of relatively short time-scale variability in the X-ray lightcurves, the results of periodic analyses of these lightcurves will now be presented. The focus in §4.4 will be on short time-scale periodicities (<500 s), while a search for longer time-scale periodicities of up to a few hours will be presented in §4.5. Some brief remarks on the long-term variability of the X-ray and optical properties will be presented in §4.6.

4.4 The 67 s X-ray periodicity

4.4.1 Search and discovery

The main period analysis tool applied in this study is the well-known Lomb-Scargle (LS) periodogram. This method involves the least-squares fitting of sinusoidal functions to data samples, and was first presented by Lomb (1976) and elaborated by Scargle (1982). The LS periodogram is a powerful timing period analysis method which is reliable not only for evenly sampled data, but also when the sampling is uneven. Horne and Baliunas (1986) further discussed the statistical properties of the LS periodogram, while Press and Rybicki (1989) presented a fast implementation of the LS periodogram that makes use of Fast Fourier Transforms (FFTs), although it is by no means equivalent to the conventional FFT periodogram analysis. The software package used for the LS analysis presented here is Starlink PERIOD⁵ version 5.0-2, which makes use of a modified version of the Press and Rybicki (1989) algorithm.

Before embarking on the period analysis, the EPIC lightcurves were detrended, i.e. the variation on time-scales longer than the length of the lightcurve were removed to prevent the build-up of significant power at low frequencies. It was found that subtracting a second-order polynomial fit from the data removed the large-scale variations quite efficiently. Consequently, a LS periodogram was created with

⁵<http://www.starlink.rl.ac.uk/docs/sun167.htx/node16.html>

PERIOD in the 2 to 100 mHz (500 s down to 10 s) range for each EPIC lightcurve with 5 s binning. The significance levels within each periodogram with respect to Gaussian white noise was determined by running a Monte Carlo significance test. The resulting periodograms for the broadband EPIC pn and MOS lightcurves of CAL 83 are presented in Fig. 4.8 and 4.9 respectively, and the broadband pn and MOS background periodograms analysed by an identical method, in Fig. 4.10 and 4.11. No periodicities were detected in the 50 to 100 mHz range, therefore only the range between 2 and 50 mHz is shown.

Upon inspection of the EPIC pn periodograms, a very strong peak at ~ 15 mHz is immediately evident in observation 0506531501, corresponding to a period of ~ 67 s. This peak is also present at a $>99.73\%$ significance level in observations 0500860601 and 0506531701, and at a lower significance level in several of the other observations. The LS periodograms of the much lower signal-to-noise MOS lightcurves did not exhibit this peak at a $>99.73\%$ level, although it does seem to be slightly above the local noise level in observations 0123510101A, 0123510101B, 0500860301, 0500860401 and especially in 0506531701. No significant peak was detected in the background periodograms of either pn or MOS. A similar analysis was performed for all the HR lightcurves, but no ~ 15 mHz peak was found. None of the periodograms exhibited harmonics of this frequency at e.g. 30 mHz or 45 mHz.

The ~ 67 s period was first discovered near the end of the M.Sc. research of the author, and preliminary results and remarks were provided in the M.Sc. dissertation (Odendaal, 2012). However, as part of this Ph.D. thesis, the properties of this ~ 67 s periodicity have now been characterized in much more detail, as presented in the rest of this section, and have also been published in MNRAS (see Appendix F).

4.4.2 Evaluating the overall significance of the ~ 67 s period

The periodograms presented in Fig. 4.8, 4.9, 4.10 and 4.11 were created by using a default Fourier frequency spacing of $1/(4T)$, where T is the total length of the particular lightcurve. In order to investigate the overall significance of the periodicity at ~ 15 mHz, all the periodograms mentioned were recreated with a constant fixed frequency interval of 0.05 mHz. For both EPIC pn and MOS, these periodograms were consequently averaged over *all 21 of the on-state datasets* (i.e. regardless of the visible presence of a 15 mHz peak), to obtain three averaged periodograms for each detector: one for the broadband CAL 83 count rate, one for the broadband background count rate and one for the CAL 83 hardness ratio. These averaged periodograms are shown in Fig. 4.12.

The strong peak at 15 mHz (66.7 s) is clearly visible far above the noise level in the

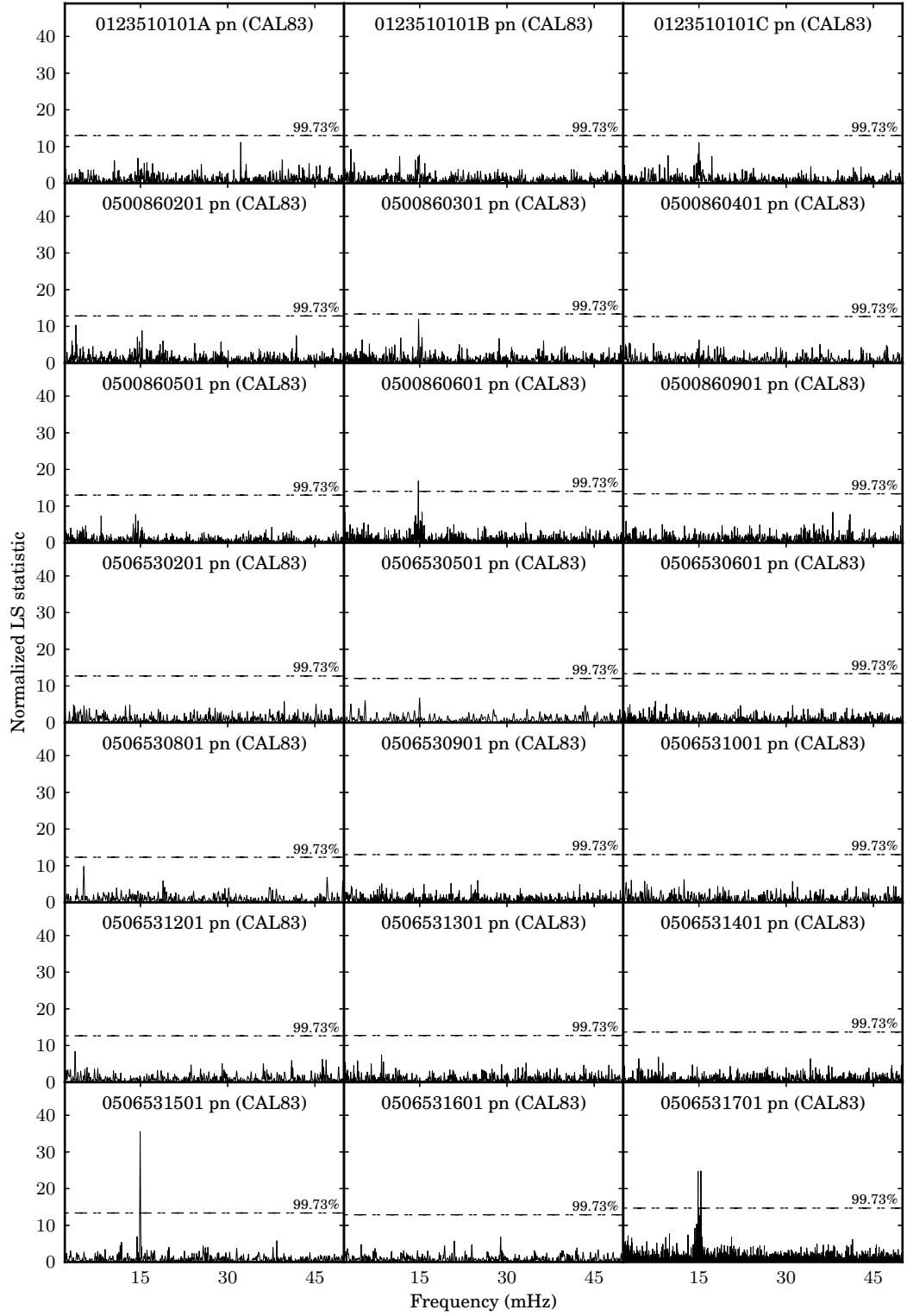


Fig. 4.8: Lomb-Scargle periodograms of the broadband EPIC pn lightcurves of CAL 83 in the 2-50 mHz range. A lightcurve binning of 5 s was used. The 99.73% significance level is indicated.

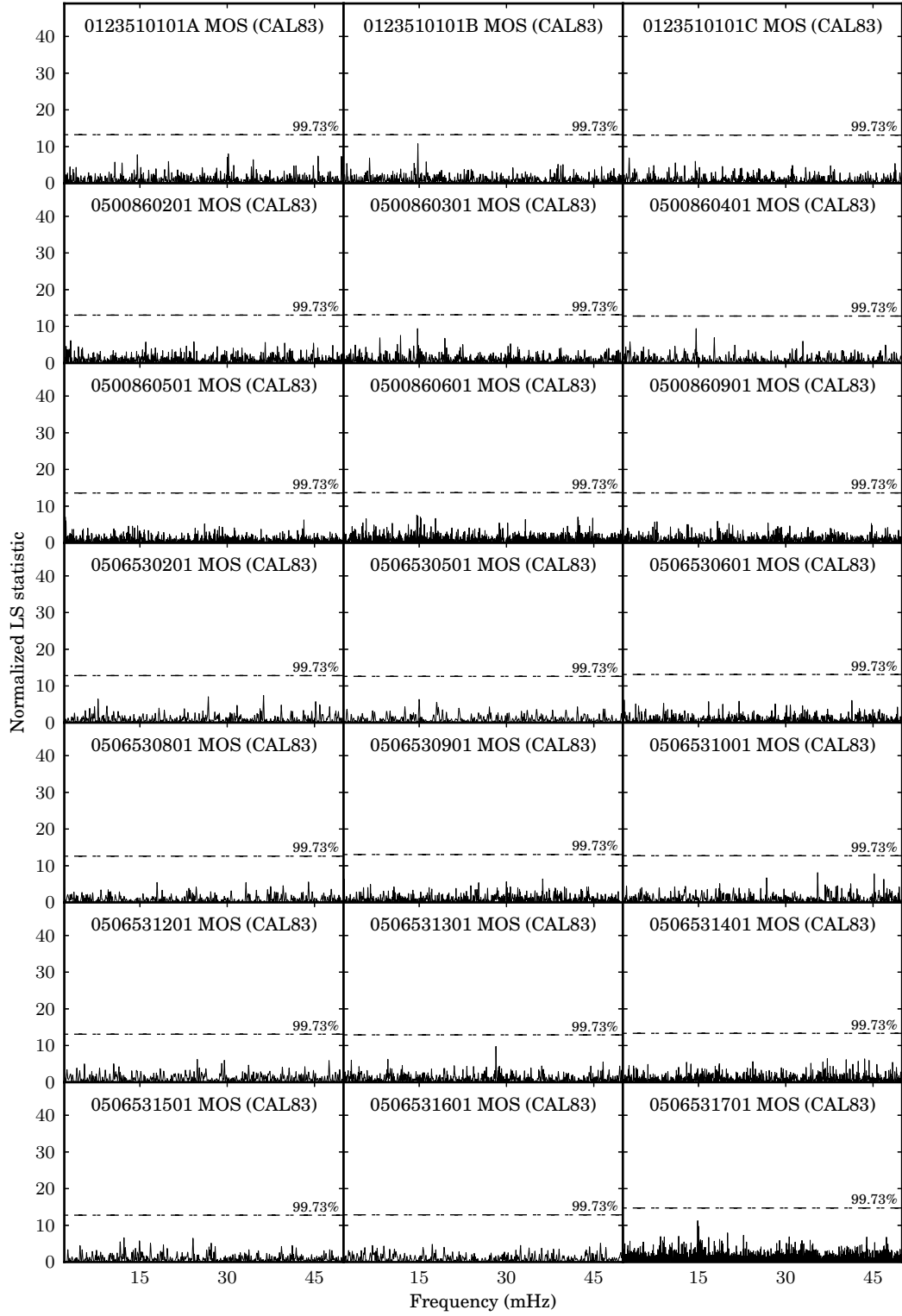


Fig. 4.9: Lomb-Scargle periodograms of the broadband EPIC MOS lightcurves of CAL 83 in the 2-50 mHz range. A lightcurve binning of 5 s was used. The 99.73% significance level is indicated.

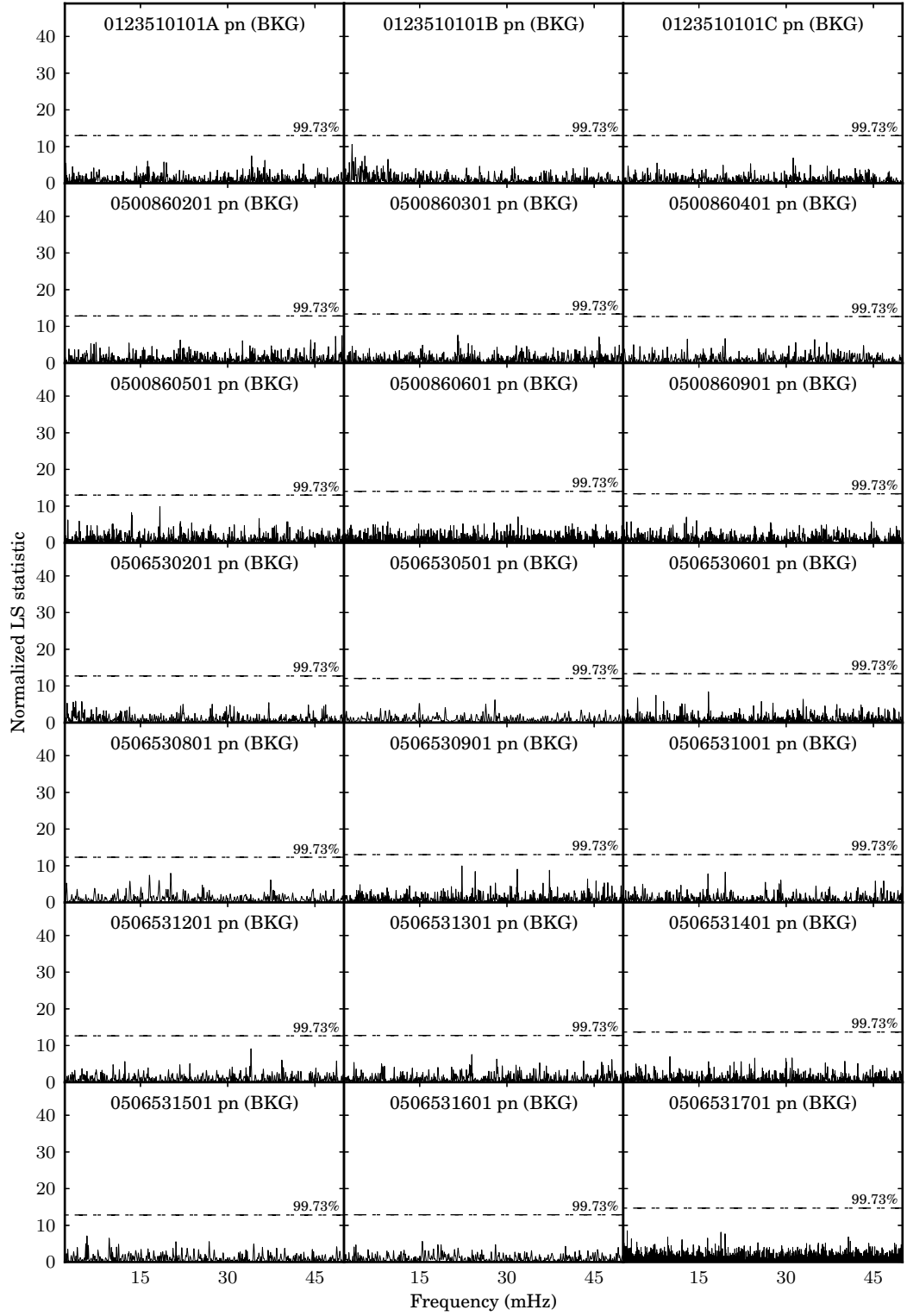


Fig. 4.10: Lomb-Scargle periodograms of the broadband EPIC pn background lightcurves in the 2-50 mHz range. A lightcurve binning of 5 s was used. The 99.73% significance level is indicated. Note the absence of significant peaks at ~ 15 mHz.

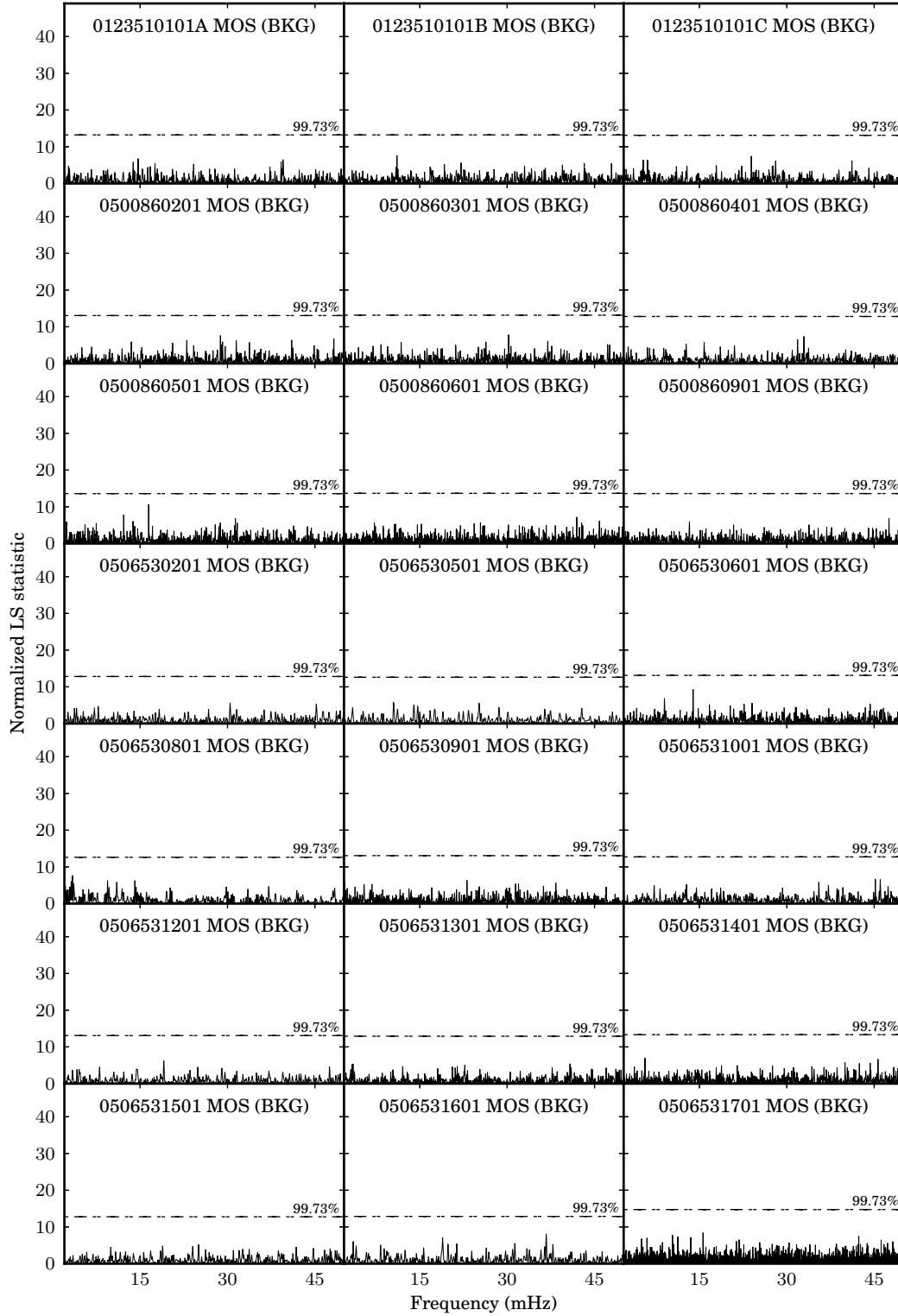


Fig. 4.11: Lomb-Scargle periodograms of the broadband EPIC MOS background lightcurves in the 2-50 mHz range. A lightcurve binning of 5 s was used. The 99.73% significance level is indicated. Note the absence of significant peaks at ~ 15 mHz.

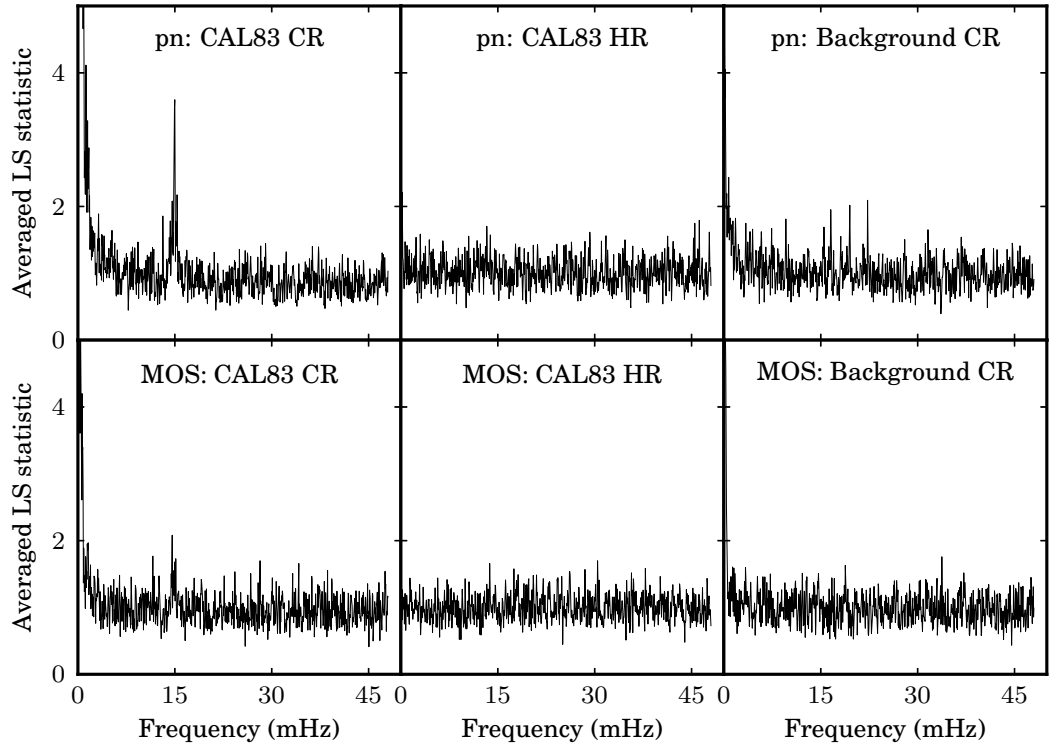


Fig. 4.12: Averaged Lomb-Scargle periodograms of EPIC broadband count rate (CR) and hardness ratio (HR) of CAL 83, and also the EPIC background, in the 0-50 mHz frequency range. A lightcurve binning of 5 s was used, with a frequency spacing of 0.05 mHz. The averaged power of the 15 mHz (67 s) periodicity is highly significant ($>3\sigma$) in the pn CR periodogram, and also slightly above the noise level in the MOS CR periodogram. Note the absence of significant peaks at ~ 15 mHz in the background and also in the HR.

average pn periodogram for CAL 83, and slightly above the noise level in the average MOS periodogram. The peak is absent in the average background periodograms, and also in the average HR periodograms. To estimate its statistical significance, the standard deviation (σ) of the 66.67 s EPIC pn peak relative to the average white noise level in the 20-50 mHz band was determined, and was found to be at a $\geq 3.5\sigma$ level.

Possible contamination from the accreting pulsar XMMU J054134.7-682550, which is situated ~ 4 arcmin away from CAL 83, was also considered. Manousakis et al. (2009) found pulsed emission from this source at a period of ~ 61 s in *XMM-Newton* observations 0500860301, 0500860401, 0500860501 and 0500860601, during which the source was observed in outburst. These authors also reported that the detected spin period of XMMU J054134.7-682550 decreased by 1.54 s during ~ 50 days as a result of spin-up effects. By careful investigation we have excluded the possibility that the ~ 67 s oscillation arises from contamination by XMMU J054134.7-682550 photons, because (i) the power spectra of this pulsar calculated from the observations under consideration do not exhibit a peak corresponding to ~ 67 s, (ii) the

~ 61 s period (which is decreasing due to spin-up effects) reported for the pulsar is easily resolvable from our ~ 67 s peaks, and (iii) the coordinates of the pulsar are significantly displaced relative to that of CAL 83 ($\sim 4'$).

After the above investigation, it was concluded that the ~ 67 s periodicity is inherent to CAL 83. Further investigation into the variability and significance of the period will now be presented. By simply scrutinizing the EPIC pn and MOS periodograms in Fig. 4.8 by eye and comparing the signal strength at 15 mHz with the local noise level, it was concluded that the ~ 67 s periodicity can be considered to be “present” in observations 0123510101 (all 3 segments), 0500860201, 0500860301, 0500860401, 0500860501, 0500860601, 0506530501, 0506531501 and 0506531701, while it is “absent” in the rest of the observations, or, perhaps more correctly, just not detectable.

Considering the mean EPIC pn count rate per observation as given in Table 4.4, one can broadly divide the *XMM-Newton* observations into two categories (apart from the four X-ray off-states): an X-ray bright state with $\text{CR} > 5 \text{ counts s}^{-1}$, and an X-ray faint state with $\text{CR} < 2 \text{ counts s}^{-1}$. The ~ 67 s periodicity was detected in all the X-ray bright observations, with an additional detection in the brightest of the faint state observations: 0506530501, with $\text{CR} = 1.680 \pm 0.205$. It is noted that this is probably a selection bias rather than an intrinsic source property, as one is more likely to detect a short time-scale variation in a lightcurve with higher counts and therefore better statistics. A detector with a larger effective area might have revealed the same periodicity in the faint observations.

4.4.3 The period and variability of the modulation

Firstly, the period value of the ~ 15 mHz (~ 67 s) peak in the periodograms and its significance level according to the Monte Carlo significance test were determined for each of the 11 datasets exhibiting this periodicity, and are presented in Table 4.7. The error estimate in the period value was obtained from the simple relation

$$\Delta P = \frac{P^2}{2T}, \quad (4.18)$$

which represents the Fourier resolution of the periodogram, where P is the value of the period itself, and T is the total length of the dataset.

Upon obtaining the period values, each of the 11 EPIC pn lightcurves was folded on the corresponding period as given in Table 4.7, using the starting point of the lightcurve as the reference point. As an additional investigation of the possible occurrence of the same periodicity in the MOS data, each MOS lightcurve was folded on the same period as the corresponding pn observation, using the same reference

Table 4.7: The exact period and significance of the ~ 67 s periodicity in the observations where it was detected, even if only at a marginal level. Values are based on the broadband EPIC pn periodograms (see Fig. 4.8). The modulation semi-amplitude was obtained by expressing the semi-amplitude of the sine fit to the folded light curve as a percentage of the mean value of the fit.

Observation ID	Period (s)	Significance (%)	Modulation semi-amplitude (%)
0123510101A	68.69 ± 0.24	29.94	2.5
0123510101B	66.86 ± 0.22	59.47	2.8
0123510101C	66.63 ± 0.22	98.16	5.0
0500860201	65.45 ± 0.19	81.69	3.1
0500860301	67.55 ± 0.22	99.02	4.0
0500860401	66.55 ± 0.30	23.25	3.9
0500860501	70.55 ± 0.20	51.74	4.0
0500860601	67.67 ± 0.11	99.99	3.7
0506530501	66.76 ± 0.49	56.03	8.6
0506531501	66.87 ± 0.35	>99.99	7.6
0506531701	65.173 ± 0.047	>99.99	2.5

point as the pn lightcurves so that their phasing can be compared. The folded lightcurves are presented in Fig. 4.13 and 4.14.

For the 11 datasets exhibiting the periodicity, a sine curve was also fitted to the data, keeping the period fixed to the value in Table 4.7. The semi-amplitude of the modulation was calculated by expressing the semi-amplitude of the sine fit to the pn data as a fraction of the mean of the fit, and is also given in Table 4.7. One can see that most of these folded lightcurves do indeed follow a quasi-sinusoidal modulation pattern, with the semi-amplitude ranging from 2.5 to 8.6%. For the 11 datasets containing the ~ 67 s periodicity, the folded pn and MOS lightcurves are in phase. The exception is 0123510101B, where the MOS lightcurve was very noisy. It should be emphasized that, although the discussion above focussed on the lightcurves with a binning of 5 s detrended with a second-order polynomial, similar results were obtained by using different lightcurve bin sizes and detrending functions, also when analysing the raw lightcurves without detrending.

The other 10 datasets where the periodicity was not detected in the periodograms, were also folded, but on the mean period of 67 s, as an additional check of whether they exhibit any traces of such a modulation. No sine curves were fitted for these. In several of these datasets, a ~ 67 s modulation may indeed be present in the data, even though it is not at a very significant level. One must also keep in mind that, if such a similar modulation is inherently contained in these observations, it may well be at a period slightly different from the ~ 67 s mean, which will change the appearance of the folded lightcurve. However, the investigation into the variability of the ~ 67 s periodicity presented in the rest of this section will focus on the other 11 EPIC datasets which do have evidence for the modulation in their periodograms.

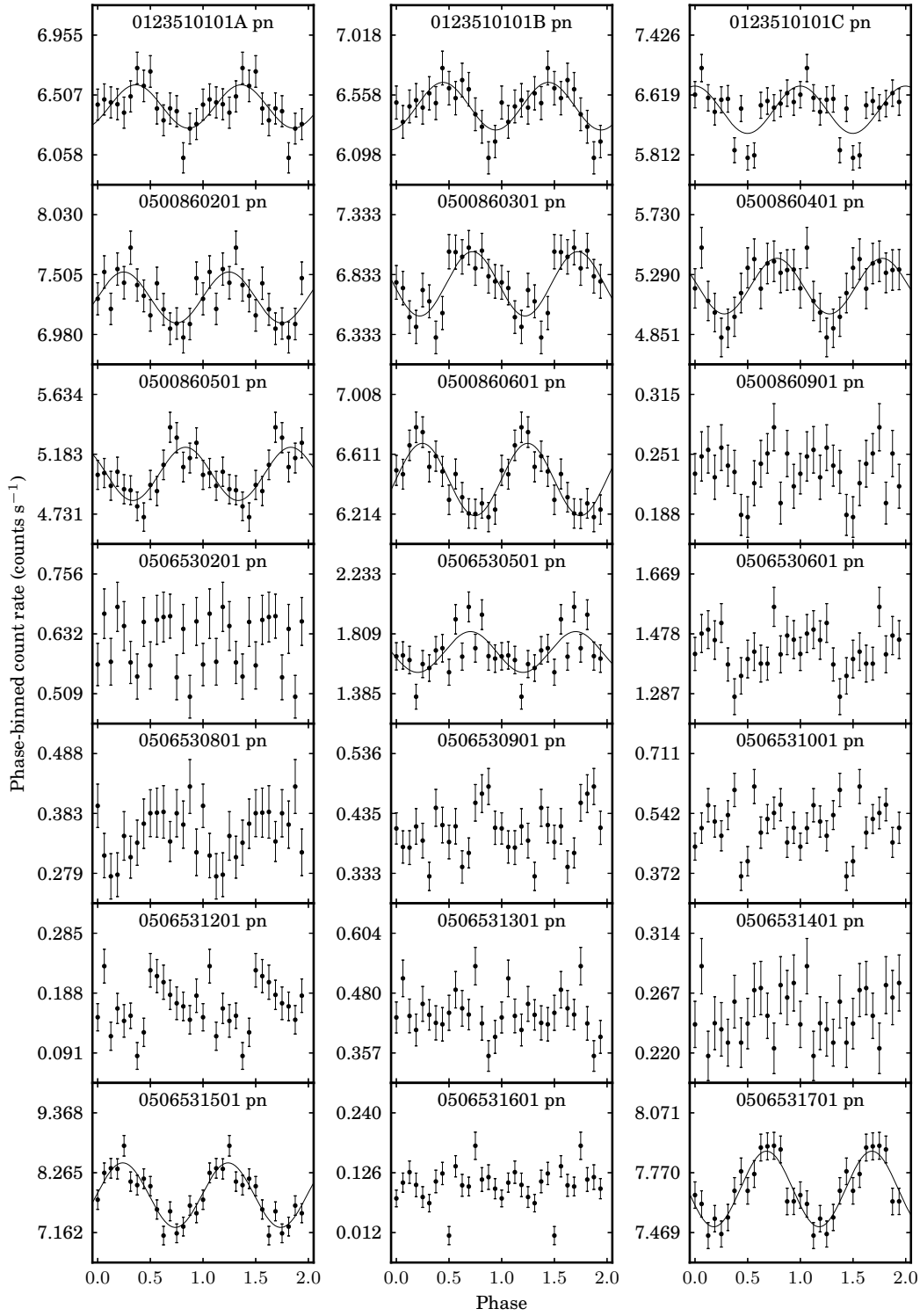


Fig. 4.13: The EPIC pn broadband lightcurves of CAL 83 folded on the ~ 67 s period. For the datasets in which the periodicity could be detected in the periodograms (see Fig. 4.8), the best fitting sine curve with the period fixed to the corresponding value provided in Table 4.7 is overplotted. The other 10 datasets were folded on the mean period of 67 s. The starting point for the phase folding was taken at the start of each dataset. Two cycles are plotted for clarity.

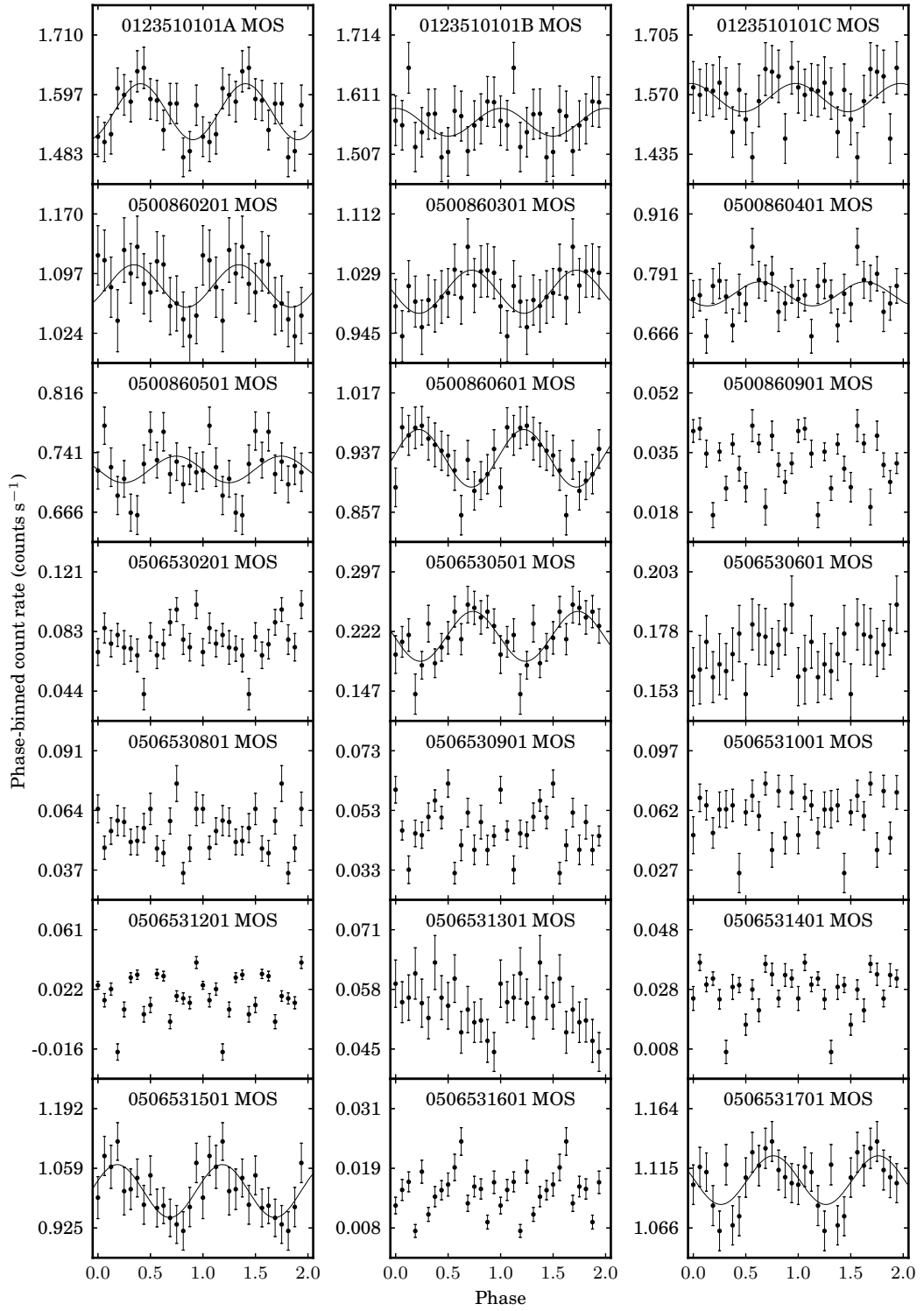


Fig. 4.14: The EPIC MOS broadband lightcurves of CAL 83 folded on the ~ 67 s period. For the datasets in which the periodicity could be detected in the periodograms (see Fig. 4.8), the best fitting sine curve with the period fixed to the corresponding value provided in Table 4.7 is overplotted. The other 10 datasets were folded on the mean period of 67 s. The starting point for the phase folding was taken at the start of the corresponding EPIC pn dataset, to enable direct comparison with the folded lightcurves in Fig. 4.13. Two cycles are plotted for clarity.

The EPIC pn hardness ratio (HR) lightcurves were also folded on the periods given in Table 4.7 (see Fig. 4.15). The shape of the folded HR lightcurves generally seem quite random, but in some cases a low amplitude quasi-sinusoidal modulation that is in phase with the CR modulation is faintly detectable, e.g. in 0123510101A, 0123510101B, 0506531501 and 0506531701.

Considering the results in Table 4.7, one can see that there is evidently considerable variability in both the period and the amplitude of the periodicity in different observations. Within a single observation, the peak often seems to be multiperiodic (see Fig. 4.8, especially observation 0506531701). This warrants further investigation into the variability of the modulation period within a single observation.

By using Eq. 4.18, it was determined that to obtain a ± 1 s error in a measurement of a period with a value near 67 s, the length of the lightcurve needs to be 2245 s. Each of the EPIC pn lightcurves with 5 s binning containing the periodicity, was consequently divided into a series of consecutive segments, each with a length of exactly 2245 s. The starting point of a particular segment was displaced by a value close to 600 s relative to the starting point of the previous segment, with the exact displacement determined by the total length of the observation, so that all the data points in the observation could be used. These overlapping segments enabled the calculation of a “moving average” of the period through the course of the observation. The lightcurves were detrended as described previously, and a Lomb-Scargle analysis and accompanying significance computation for frequencies between 10 and 20 mHz were performed for each segment.

The results of this analysis are shown in the informative colour images in Fig. 4.16, together with the mean EPIC pn count rate for comparison. The segments in which a period between 13.5 mHz and 16.5 mHz was found at a $>95.45\%$ level are annotated with the period and its significance. It is noted that in the observations with the lowest count rates (0500860401 and 0506530501), no segments with a $>95.45\%$ detection were found.

It is evident that there is significant variability of at least ~ 3 s to each side of the mean of the ~ 67 s periodicity. Significant variability is even present within a single dataset. In the next subsection, the possibility of a correlation between the value of the period and the mean X-ray count rate will be investigated.

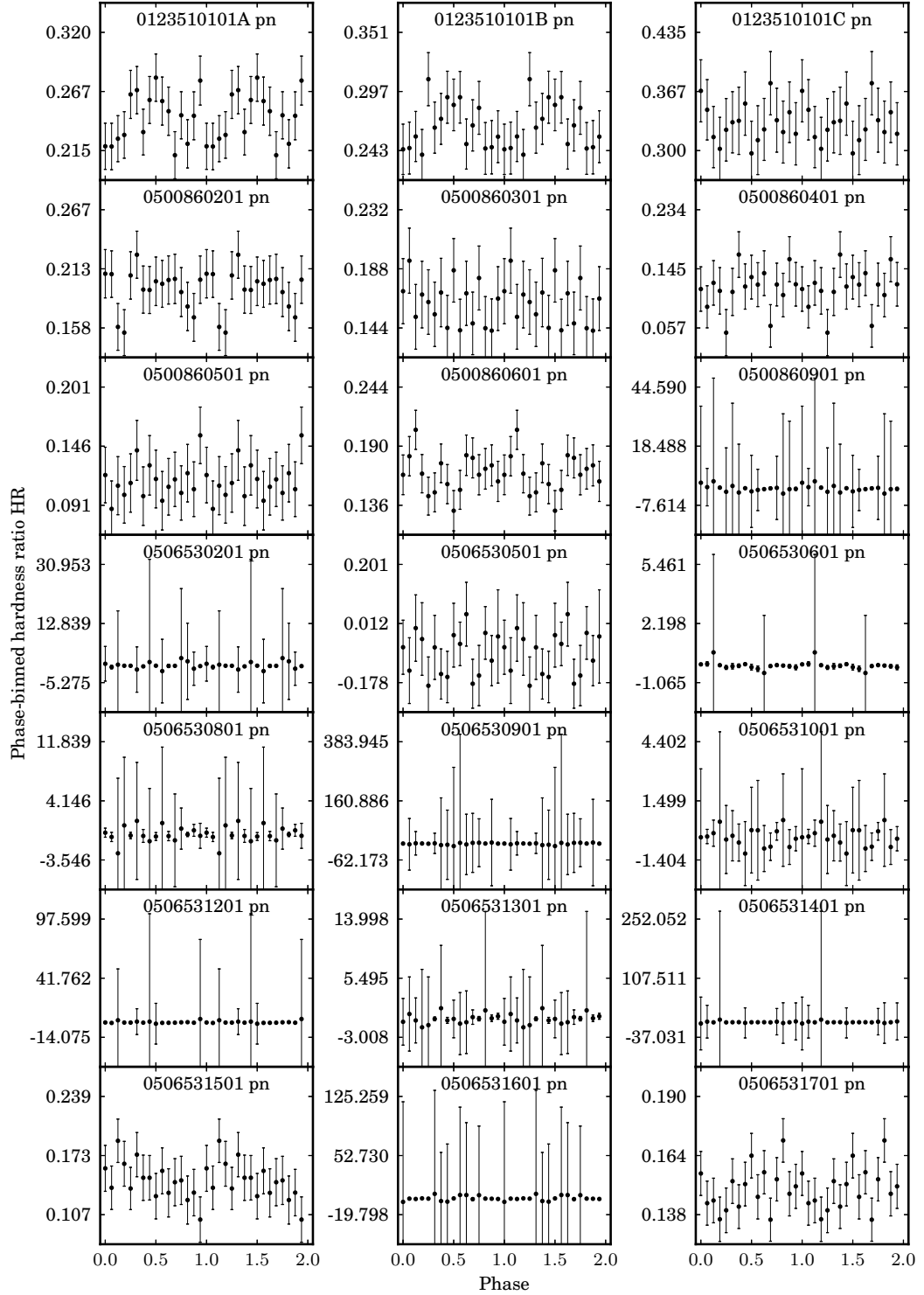


Fig. 4.15: The EPIC pn hardness ratio (HR) lightcurves of CAL 83 folded on the ~ 67 s period. The starting point for the phase folding was taken at the start of each dataset. Two cycles are plotted for clarity.

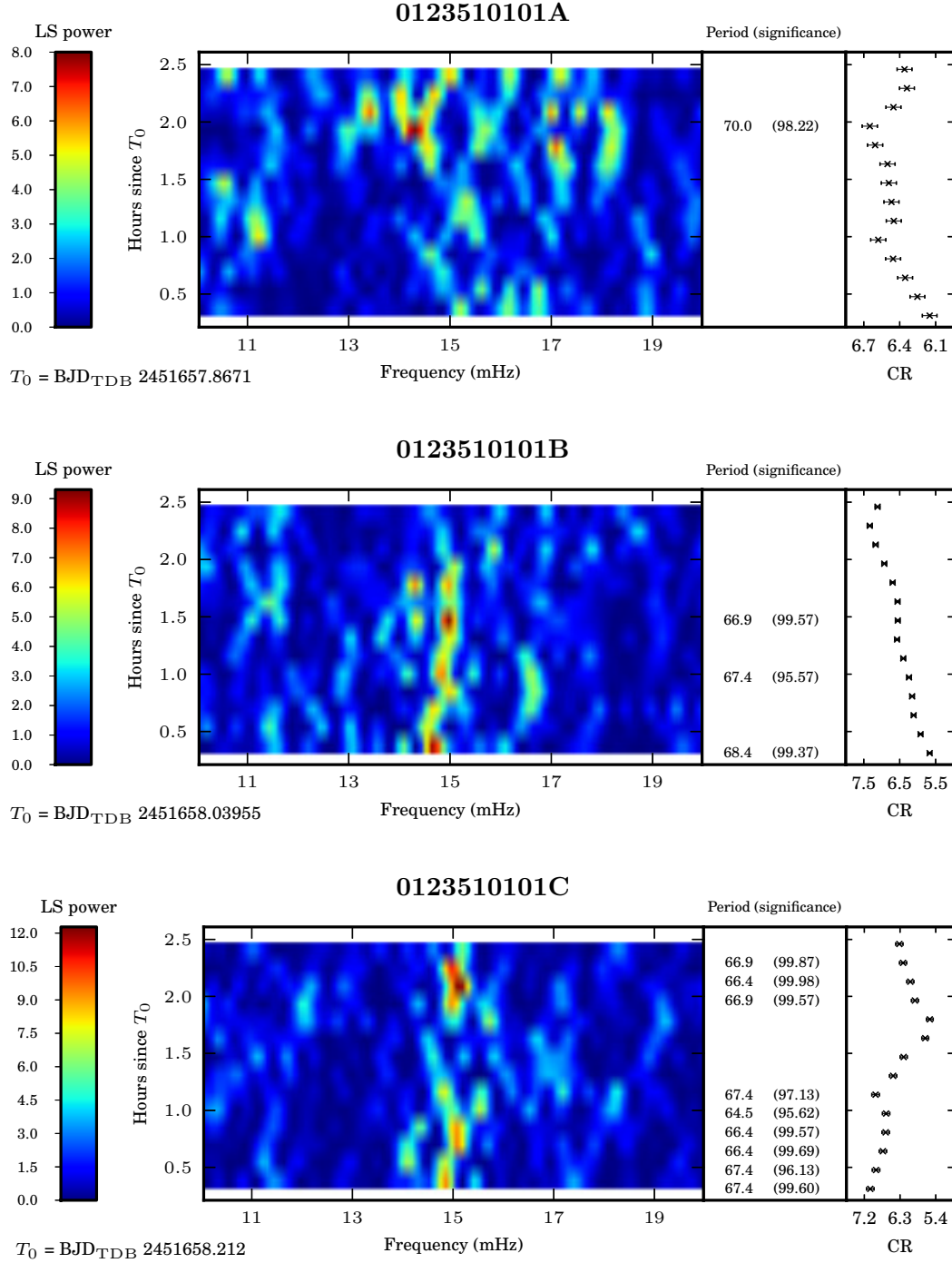


Fig. 4.16: Lomb-Scargle periodograms of EPIC pn data of CAL 83, illustrating the variability in the value of the ~ 67 s period. The colour represents the Lomb-Scargle power, and in cases where a peak was detected between 13.5 and 16.5 mHz at a $>95.45\%$ significance level, the corresponding period in seconds and its significance percentage are given. The error in the period values is ± 1 s. The BJD_{TDB} reference represents the start of the observation. For a comparison with the X-ray count rate (CR), the broadband lightcurve is also provided.

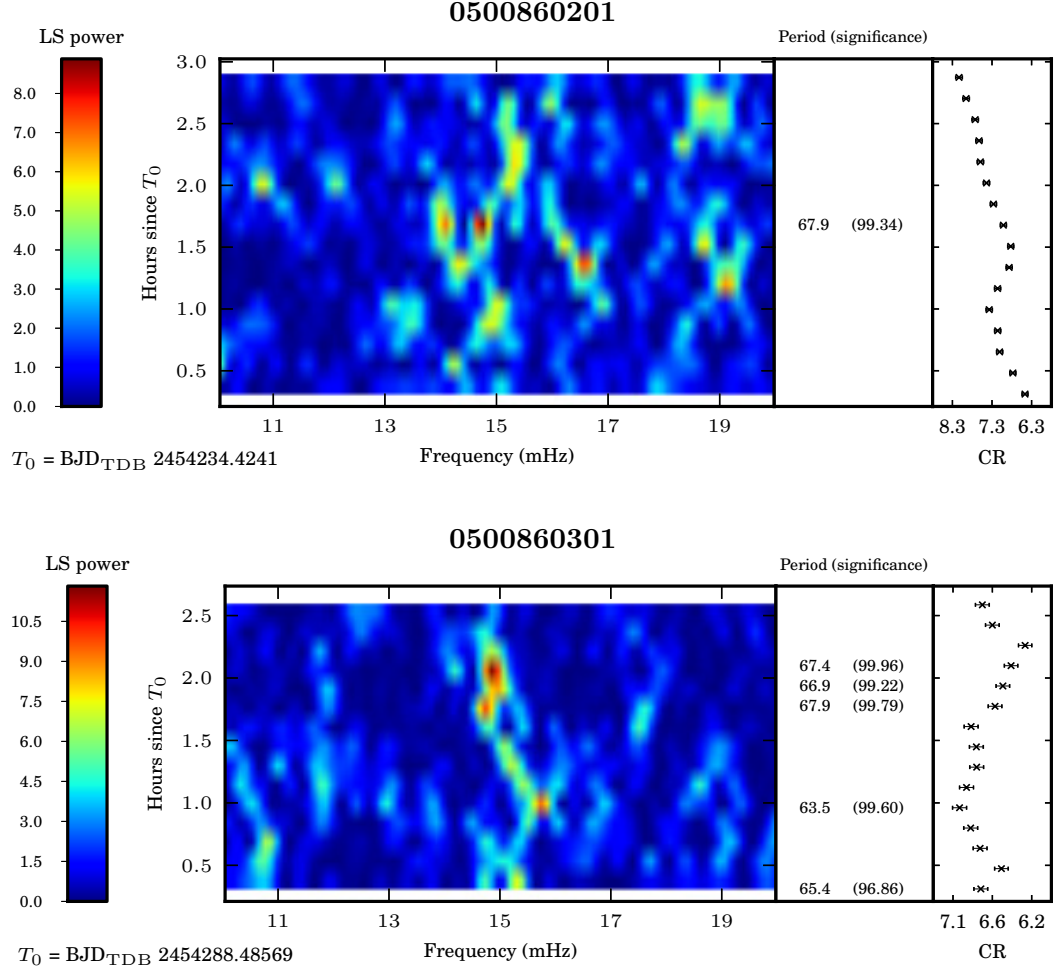


Fig. 4.16: (continued) Lomb-Scargle periodograms of EPIC pn data of CAL 83, illustrating the variability in the value of the ~ 67 s period. The colour represents the Lomb-Scargle power, and in cases where a peak was detected between 13.5 and 16.5 mHz at a $>95.45\%$ significance level, the corresponding period in seconds and its significance percentage are given. The error in the period values is ± 1 s. The BJD_{TDB} reference represents the start of the observation. For a comparison with the X-ray count rate (CR), the broadband lightcurve is also provided.

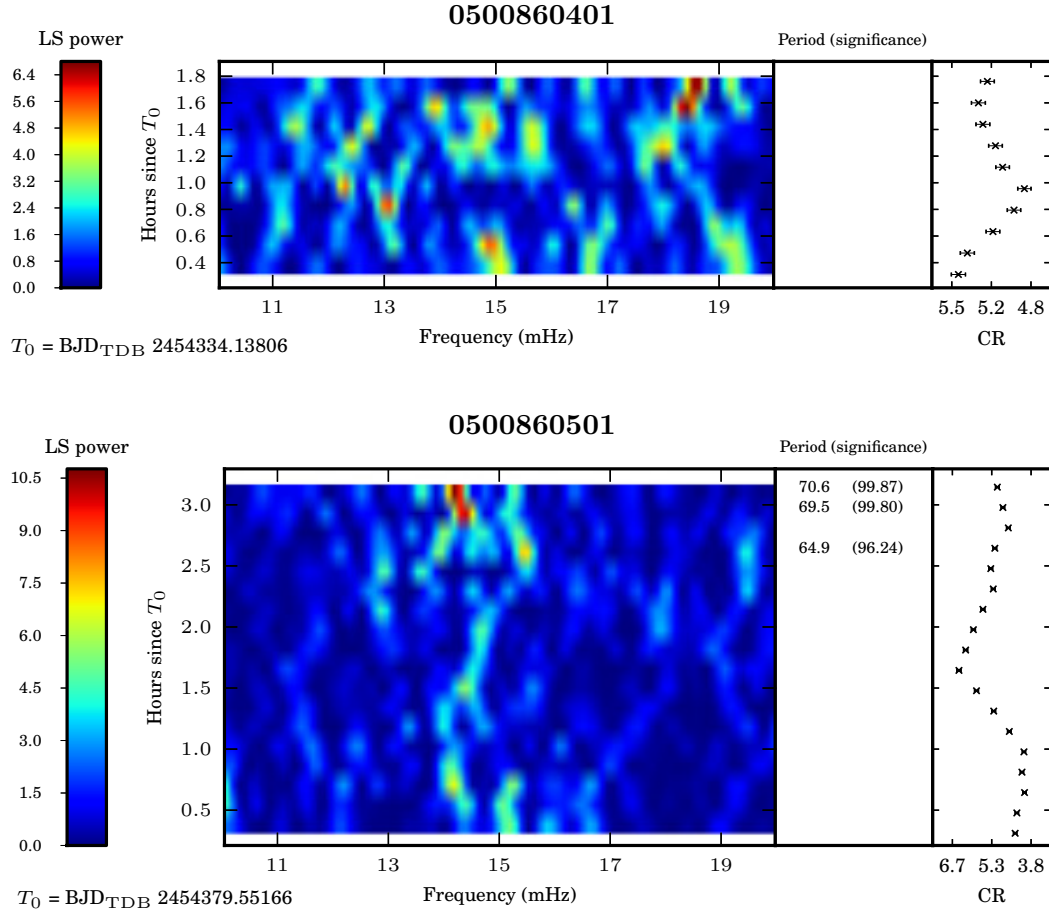


Fig. 4.16: (continued) Lomb-Scargle periodograms of EPIC pn data of CAL 83, illustrating the variability in the value of the ~ 67 s period. The colour represents the Lomb-Scargle power, and in cases where a peak was detected between 13.5 and 16.5 mHz at a $>95.45\%$ significance level, the corresponding period in seconds and its significance percentage are given. The error in the period values is ± 1 s. The BJD_{TDB} reference represents the start of the observation. For a comparison with the X-ray count rate (CR), the broadband lightcurve is also provided.

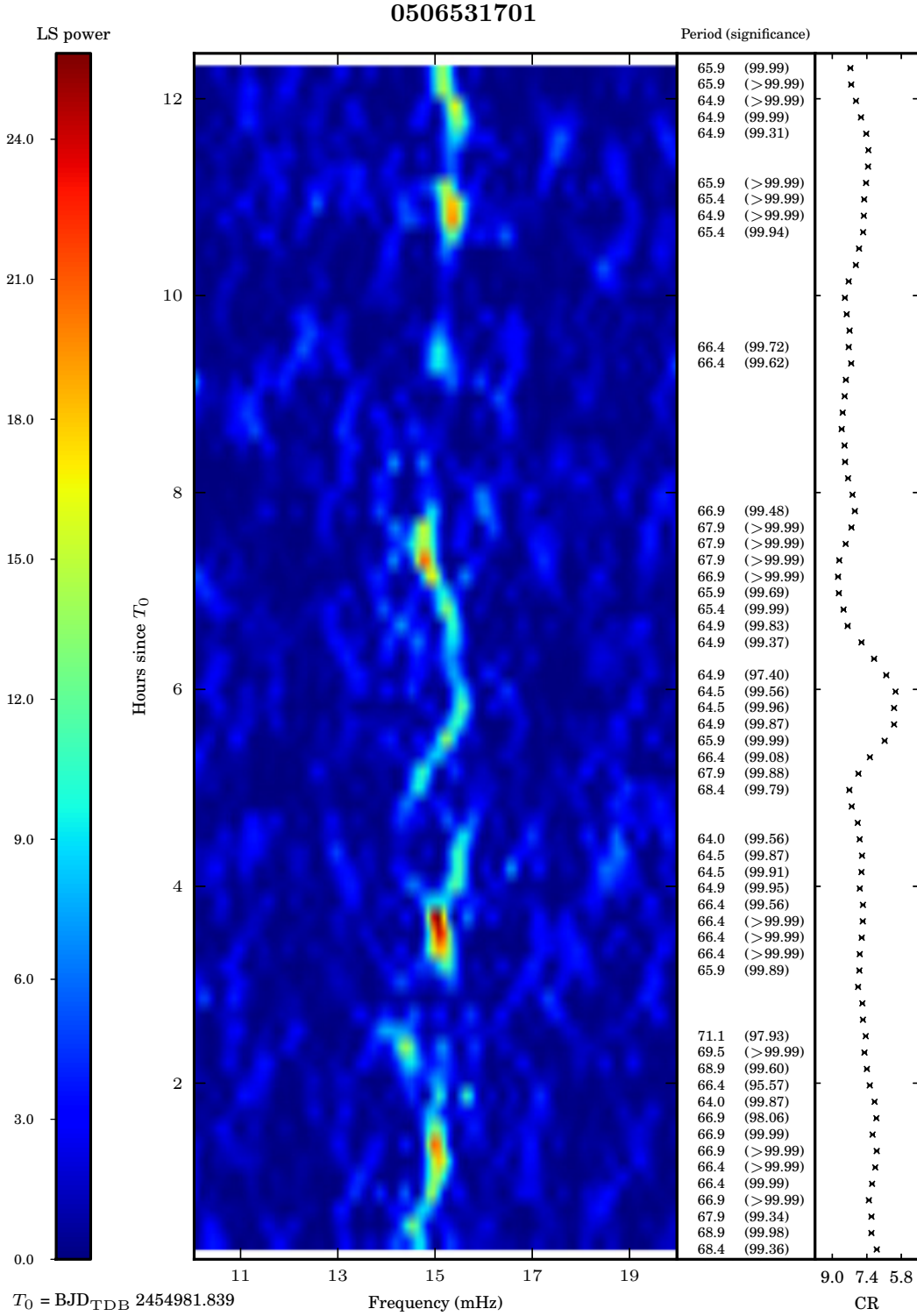


Fig. 4.16: (continued) Lomb-Scargle periodograms of EPIC pn data of CAL 83, illustrating the variability in the value of the ~ 67 s period. The colour represents the Lomb-Scargle power, and in cases where a peak was detected between 13.5 and 16.5 mHz at a $>95.45\%$ significance level, the corresponding period in seconds and its significance percentage are given. The error in the period values is ± 1 s. The BJD_{TDB} reference represents the start of the observation. For a comparison with the X-ray count rate (CR), the broadband lightcurve is also provided.

4.4.4 Testing the correlation between period and X-ray count rate

A Spearman linear rank-order correlation test (see §4.3.3) was performed by comparing the mean period value in the observation as a whole (Table 4.7) to the mean EPIC pn count rate (CR) (as given in Table 4.4). The correlation coefficient is -0.373 , with a significance value of 74.11%. This indicates the possibility of an anti-correlation of the period and the CR, i.e. lower periods being associated with higher CRs, although the significance is not very high.

Hence, the trends in the period and CR in the different segments *within* the observations were investigated by scrutinizing Fig. 4.16. Datasets 0123510101B and 0500860301 suggest the possibility of an anti-correlation between period value and CR, while 0506531701 and perhaps 0500860601 seem to rather exhibit a correlation. In the increase at ~ 0.5 h since T_0 in observation 0506531701, it seems as if the period modulation might be slightly lagging behind the CR modulation.

Datasets exhibiting an overall significance $>95.45\%$ of the ~ 67 s peak for the whole observation were 0123510101C, 0500860301, 0500860601, 0506531501 and 0506531701 (Table 4.7). Datasets 0123510101C and 0506531501 are quite short, as is 0500860301, in which the period was also only detected in a few segments (see Fig. 4.16). Observations 0500860601 and 0506531701 are much longer, and therefore the most suited to investigate the correlation between the period value and the CR.

Consequently, the period and CR values corresponding to a $>95.45\%$ detection were extracted from the 0500860601 and 0506531701 data in Fig. 4.16, and a Spearman linear rank-order correlation calculated for both sets. The correlation coefficient and its significance is $+0.152$ and 43.96% for 0500860601, and $+0.080$ and 42.45% for 0506531701, which can not be considered as a significant correlation.

4.4.5 Calculating Q for the ~ 67 s period

As explained in §2.7.4, the quality factor Q can be used to quantify the rate of change of a periodicity P , and is given by $|dP/dt|^{-1}$. The more rapid the variability in P , the smaller the value of Q . The value of Q therefore provides a measure of the coherence of the periodicity. To quantify the variability in the ~ 67 s periodicity, the period values and corresponding times of observation t was extracted from the longest observations, i.e. 0500860601 and 0506531701, selecting data only from the segments in Fig. 4.16 with a $>95.45\%$ detection significance. The value of Q was evaluated for every set of two measurements (t_i, P_i) and (t_{i+1}, P_{i+1}) in the lightcurve

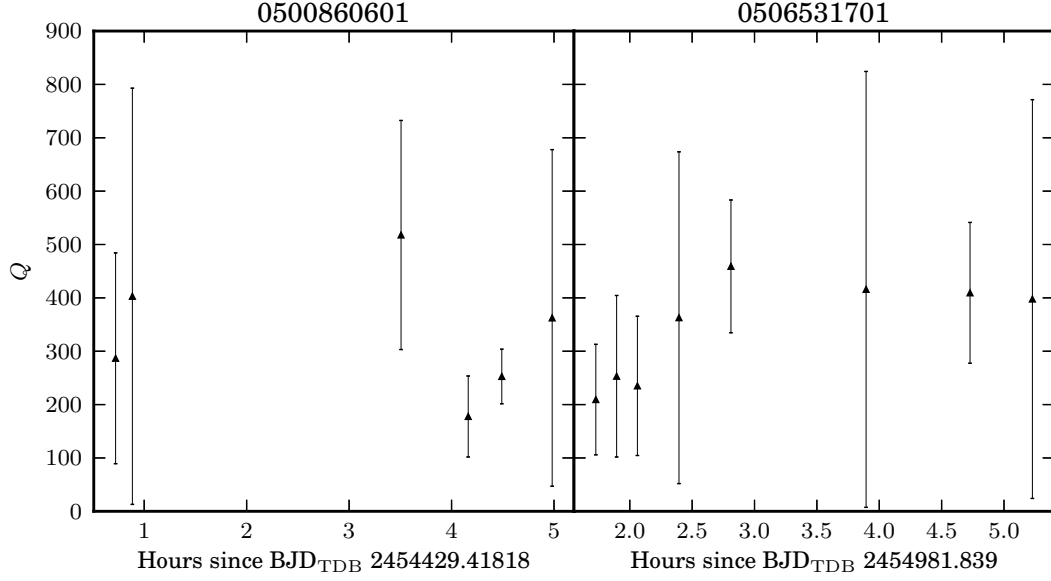


Fig. 4.17: The quality factor (Q) of the ~ 67 s period in the longest EPIC pn observations. Non-detections were not plotted, i.e. where the error of the Q -value included zero.

(provided that $P_i \neq P_{i+1}$), using

$$Q = \left| \frac{P_{i+1} - P_i}{t_{i+1} - t_i} \right|^{-1}, \quad (4.19)$$

yielding a value for Q at time $(t_i + t_{i+1})/2$. Standard propagation of the errors in the periods were used to obtain the errors in Q . The resultant values of Q as a function of time in these two observations are shown in Fig. 4.17. Only Q -values that for which the error bars did not include zero were plotted. For observation 0500860601, Q -values were between 178 ± 76 and 520 ± 210 , with a mean of 333 ± 98 , and in observation 0506531701, between 210 ± 100 and 460 ± 120 , with a mean of 343 ± 87 .

As discussed in §3.1.1, CAL 83 has been observed by several X-ray satellites apart from *XMM-Newton*. However, the effective areas of these other X-ray telescopes are 1 or more orders of magnitude smaller than that of *XMM-Newton* in the 0.15–1.0 keV band. The only on-state observation of CAL 83 with *Chandra* was obtained with the HRC-S, using the Low Energy Transmission Grating. Although *Chandra* has the highest sensitivity among these other telescopes, its effective area when using the grating is still approximately an order of magnitude smaller than that of *XMM-Newton* EPIC pn. However, the energy response of HRC-S/LETG is calibrated down to lower energies than the *XMM-Newton* detectors, making *Chandra* better suited for spectral analysis of SSSs.

As an additional test, a Lomb-Scargle analysis was performed on the *Chandra* on-state dataset, as well as all the archival *Swift* X-ray Telescope (XRT) data of CAL 83.

It was noted that the count rates were much lower than in the EPIC pn lightcurves, as expected from the sensitivity differences. The *Chandra* and *Swift* periodograms did not exhibit a significant peak at ~ 67 s. This does not necessarily mean that the periodicity was absent in the source during these observations; it may only be that the S/N of the lightcurves were too low for the periodicity to be detected.

4.5 X-ray periodicities on time-scales of several minutes

After the investigation into the short ~ 67 s period, the search was extended to longer periods, i.e. frequencies smaller than 2 mHz. As potential periodicities in this region can be a substantial fraction of the length of an observation, the lightcurves were then detrended simply by subtracting a linear fit to avoid distorting the signal. A Lomb-Scargle analysis and Monte Carlo significance tests were performed as before. The resulting periodograms for CAL 83 created from the broadband EPIC pn and MOS lightcurves with a binning of 100 s are presented in Fig. 4.18 and 4.19, and the background periodograms created similarly in Fig. 4.20 and 4.21.

Considering the pn source periodograms, peaks are visible below 1 mHz (>1000 s) in several of the datasets. As a first selection, those peaks with a significance level $\geq 99.73\%$ were selected from the periodograms. However, determining the significance of peaks in the low frequency region is notoriously difficult, due to the possible presence of red noise in this region. Consequently, the same analysis was repeated by using different lightcurve bin sizes, and also by analysing slightly shortened segments consisting of 70%, 80% and 90% of the original lightcurve, extracted from the beginning, middle and end of the original lightcurve.

Only those peaks from the initial identification list which remained stable at the same position during the various iterations, were considered to represent significant periodicities in the data. These peaks are indicated in Fig. 4.18, and the period values are given in Table 4.8. It was noted that the peak significances are higher when using smaller bin sizes, bringing some of the weaker peaks to slightly above the 99.73% threshold. This behaviour can be expected, since by increasing the binning, one is increasingly smoothing out the signal. However, to be conservative, the results obtained with 100 s binning are quoted here, but it may be possible that other weak periodicities inherent to the source may also be present in the data.

A similar analysis was performed for the hardness ratio lightcurves, but no significant peak was detected for either pn or MOS, with the exception of a peak in observation 0506531701 just touching the 99.73% level at a position corresponding to that of the strongest peak of this observation in Fig. 4.19.

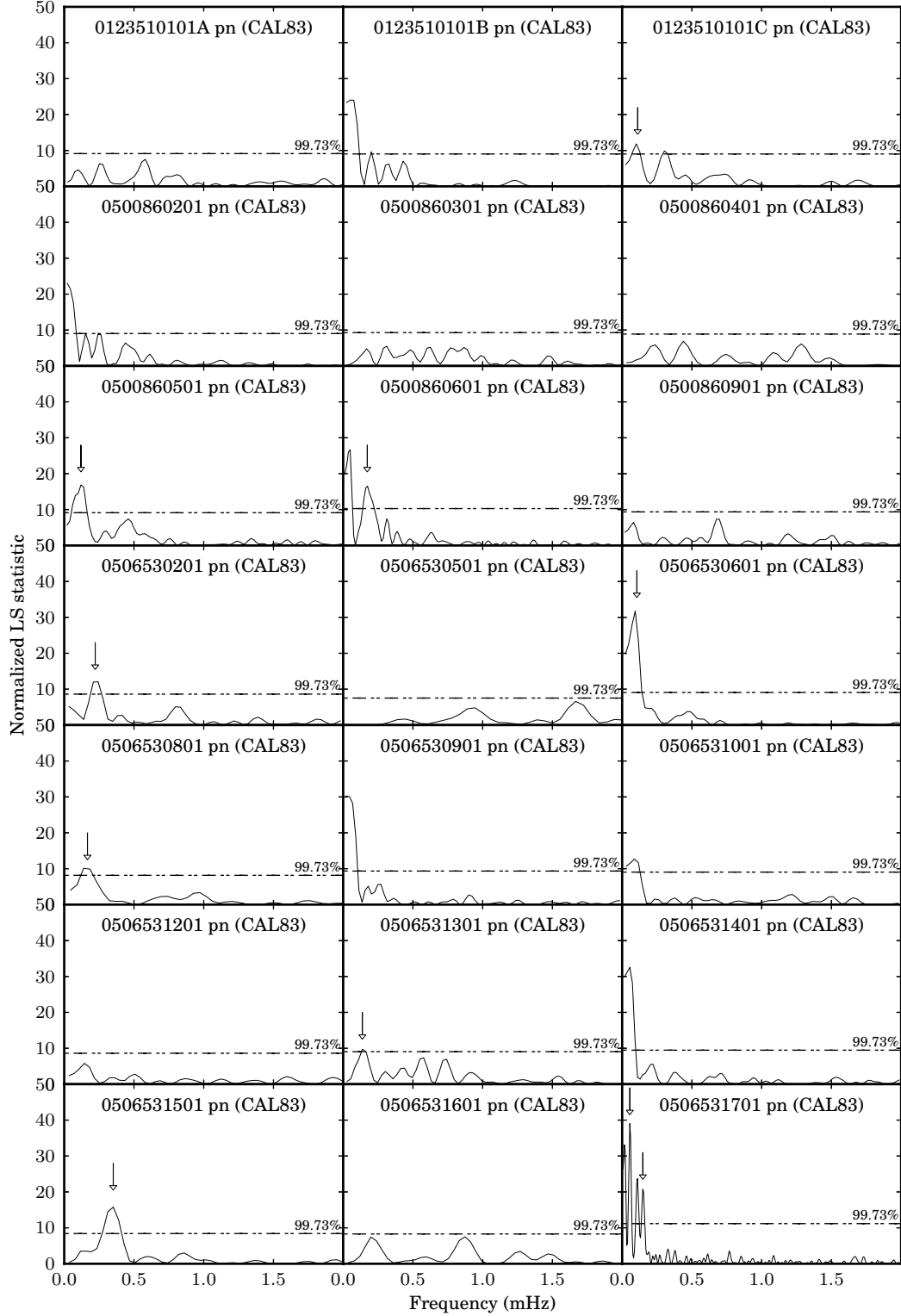


Fig. 4.18: Lomb-Scargle periodograms of the broadband EPIC pn lightcurves of CAL 83 in the 0-2 mHz range. A lightcurve binning of 100 s was used. The 99.73% significance level is indicated. The peaks considered to represent inherent source periodicities are indicated with arrows (see text).

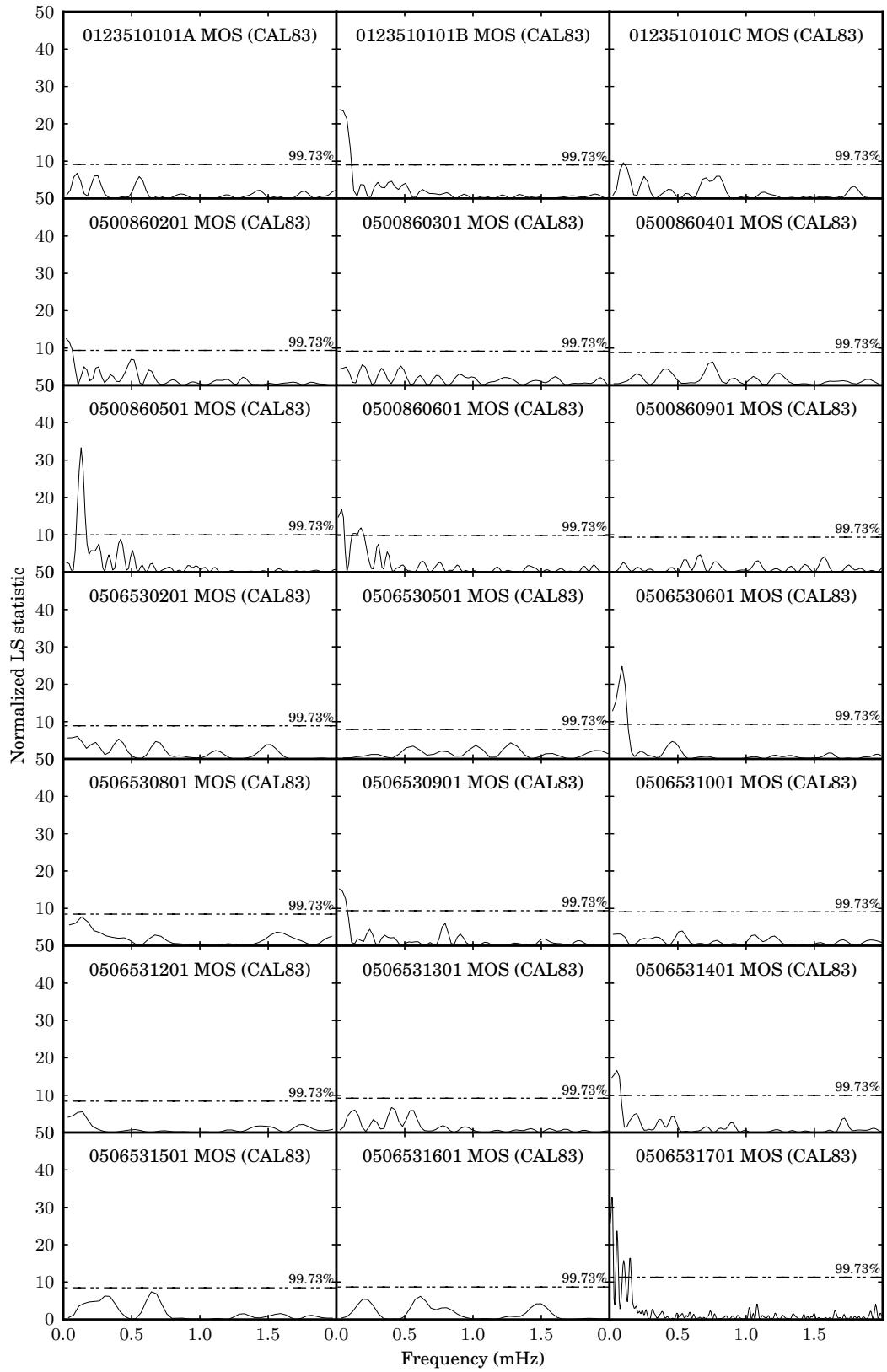


Fig. 4.19: Lomb-Scargle periodograms of the broadband EPIC MOS lightcurves of CAL 83 in the 0-2 mHz range. A lightcurve binning of 100 s was used. The 99.73% significance level is indicated.

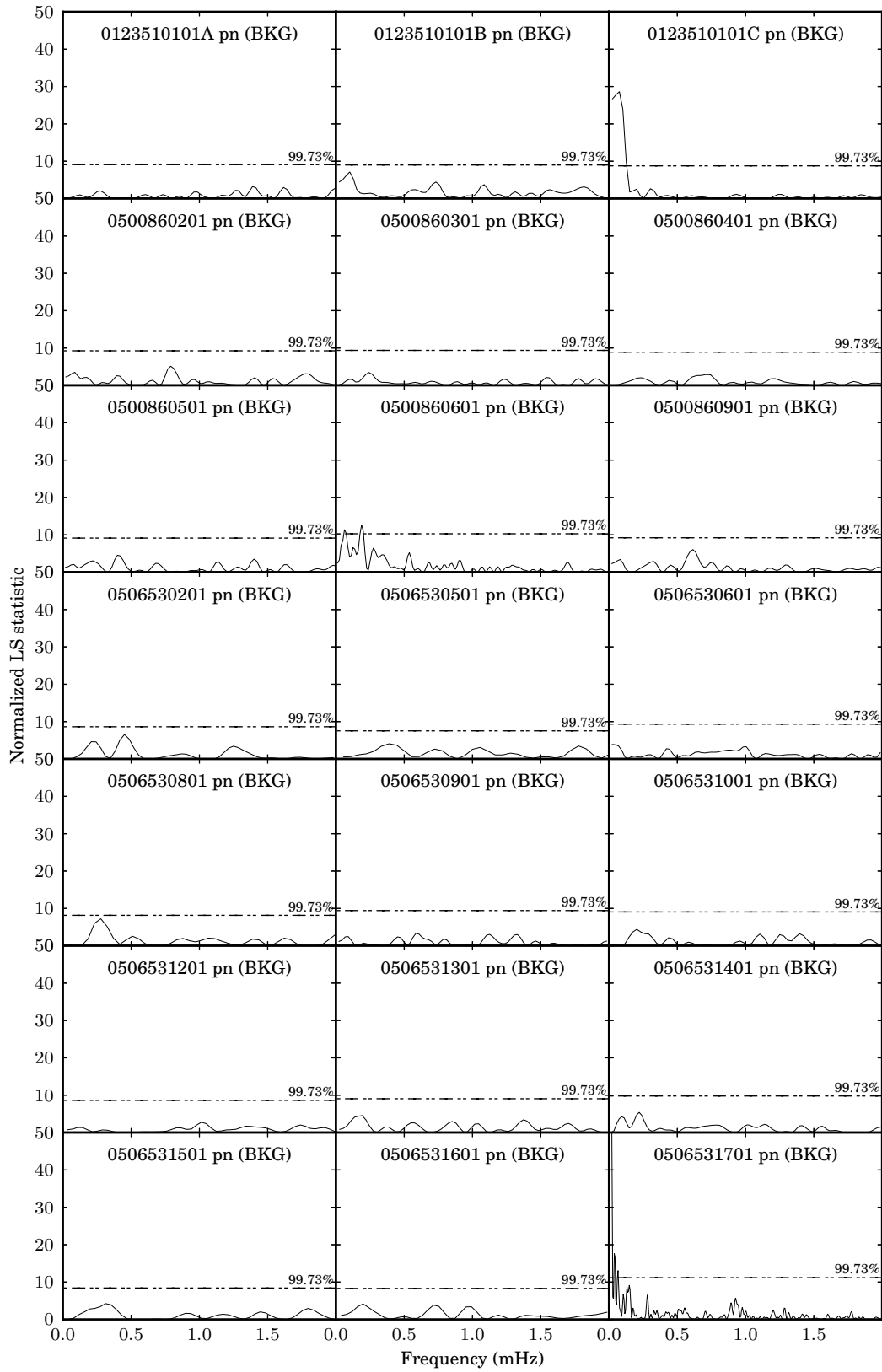


Fig. 4.20: Lomb-Scargle periodograms of the broadband EPIC pn background lightcurves in the 0-2 mHz range. A lightcurve binning of 100 s was used. The 99.73% significance level is indicated.

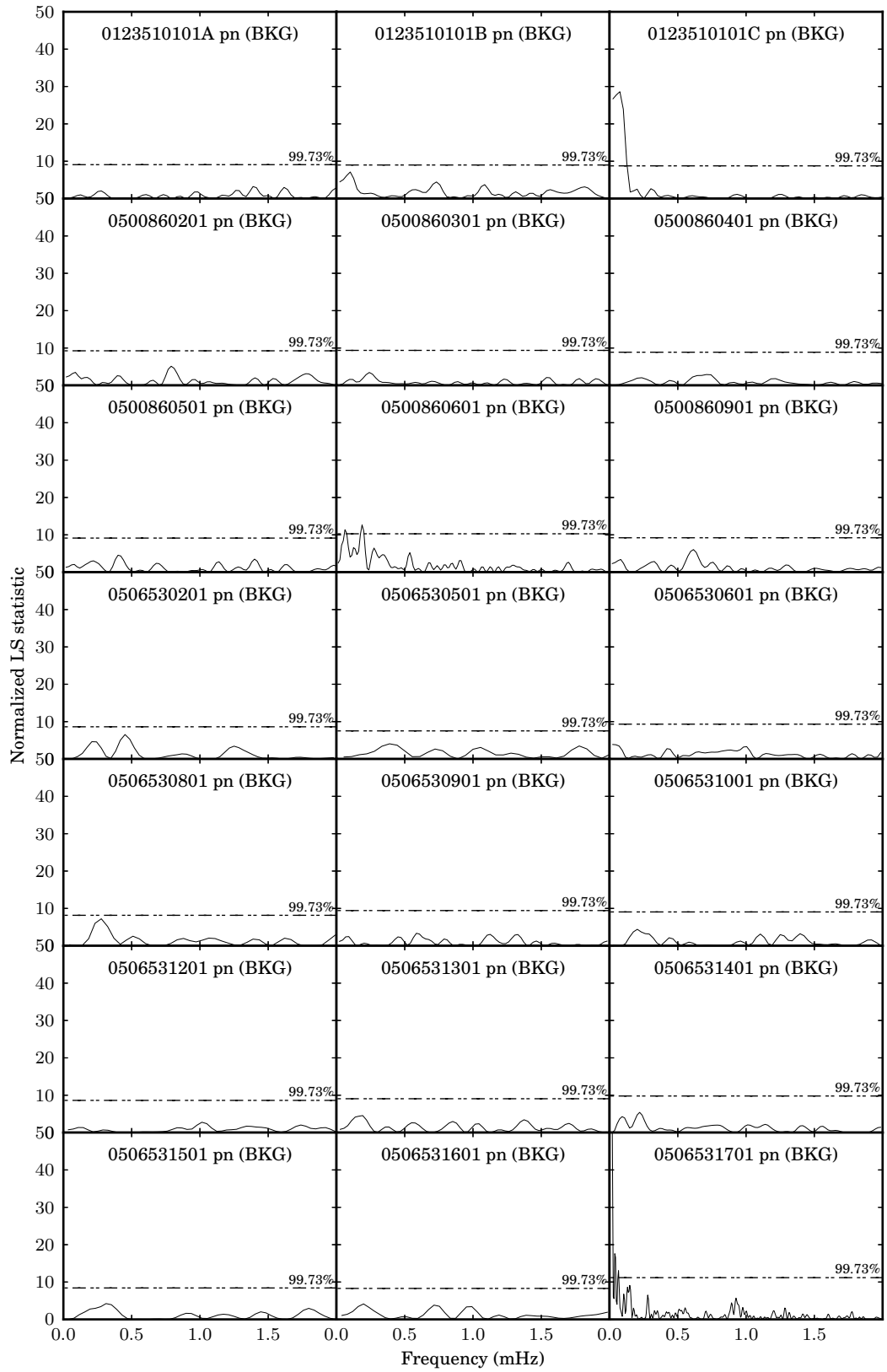


Fig. 4.21: Lomb-Scargle periodograms of the broadband EPIC MOS background lightcurves in the 0-2 mHz range. A lightcurve binning of 100 s was used. The 99.73% significance level is indicated.

Table 4.8: The period values corresponding to the power peaks detected above a 99.73% significance level over white noise in the periodograms of CAL 83 . Values are based on the broadband EPIC pn periodograms (see Fig. 4.18).

Observation ID	Period (min)
0123510101C	132 $\pm$ 53
0500860501	139 $\pm$ 46
0500860601	83 $\pm$ 10
0506530201	96 $\pm$ 38
0506530601	143 $\pm$ 57
0506530801	90 $\pm$ 45
0506531301	121 $\pm$ 49
0506531501	47 $\pm$ 11
0506531701	304 $\pm$ 61
0506531701	112.6 $\pm$ 8.3

These peaks were not detected at a significant level in the background periodograms of either pn or MOS, with the exception of observations 0500860601 and 0506531701. These peaks are much weaker in the background periodograms, therefore their presence in the source periodograms would not be only due to a background signal. It is, however, probable that contamination of the background region by the nearby source region in these observations could have introduced the faint periodicity into the background lightcurves as well.

The EPIC pn broadband lightcurves folded on these periods are shown in Fig. 4.22. In several of the folded lightcurves, the modulation pattern is quasi-sinusoidal, while in other cases it is more irregular. As an additional check for similar modulations in the hardness ratio, the corresponding HR lightcurves were also folded on these periods, as shown in Fig. 4.23. Many of the latter exhibit no regular variation, but in some of them a modulation that is in phase with the modulation in the folded broadband lightcurves is indeed visible. Therefore it does seem as if (at least some of) the longer time-scale periodicities discussed here may be accompanied by a corresponding change in the hardness ratio. This can be expected, considering the results of the correlation studies in §4.3.3.

4.6 Long-term variability

Regular photometric observations of CAL 83 have been obtained from January 1993 up to January 2000 as part of the MACHO (MASSive Compact Halo Objects) Project, and also since September 2001 as part of the OGLE (Optical Gravitational Lensing Experiment) Project. A more detailed discussion of these lightcurves will be provided in §6.2. In this section, the focus will be on the comparison of the variability of the quantities measured from the *XMM-Newton* data with the superorbital optical variability as defined by the OGLE-III lightcurve spanning the observations.

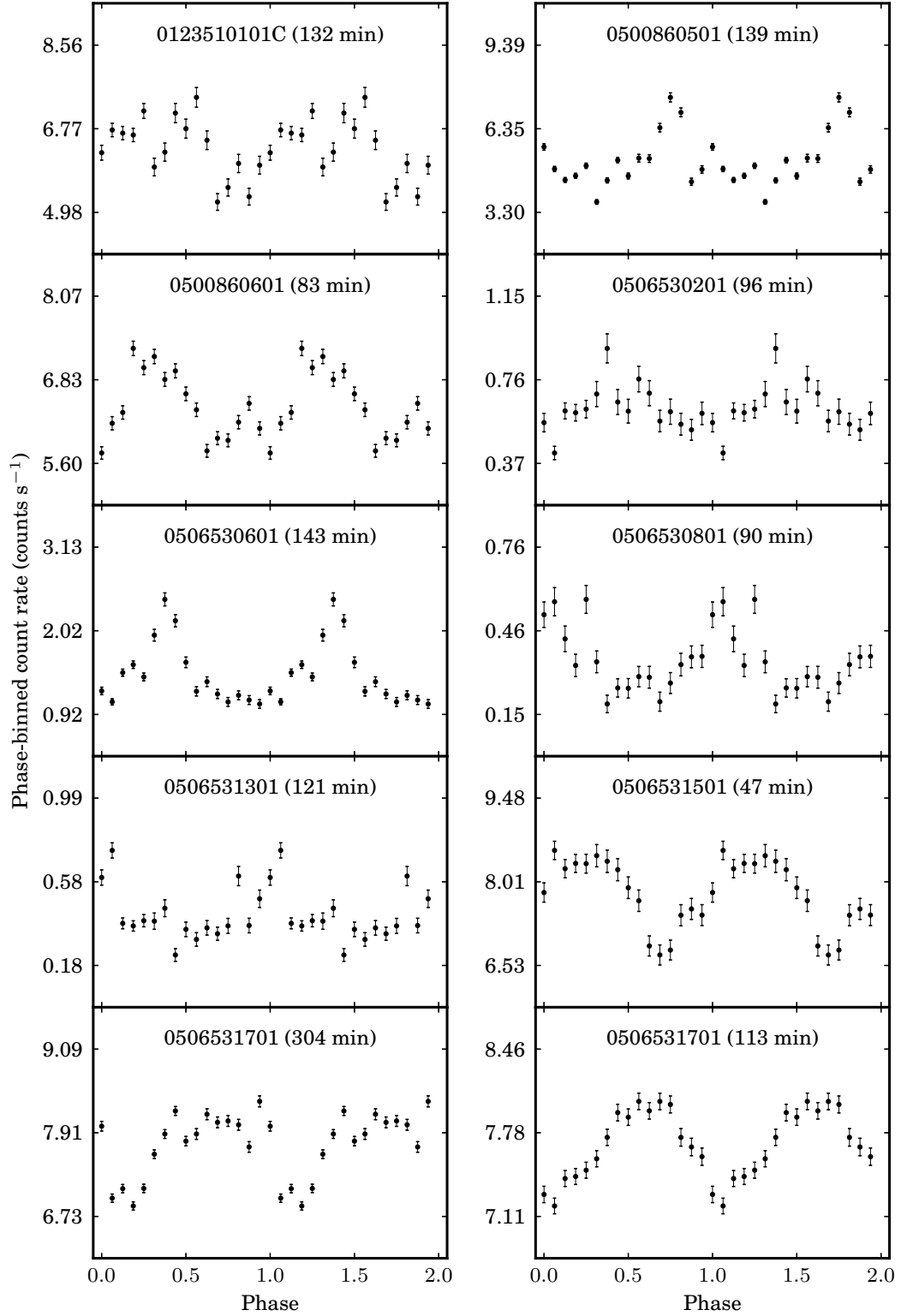


Fig. 4.22: The EPIC pn broadband lightcurves folded on the longer periods listed in Table 4.8. The starting point for the phase folding was taken at the start of each dataset. Two cycles are plotted for clarity.

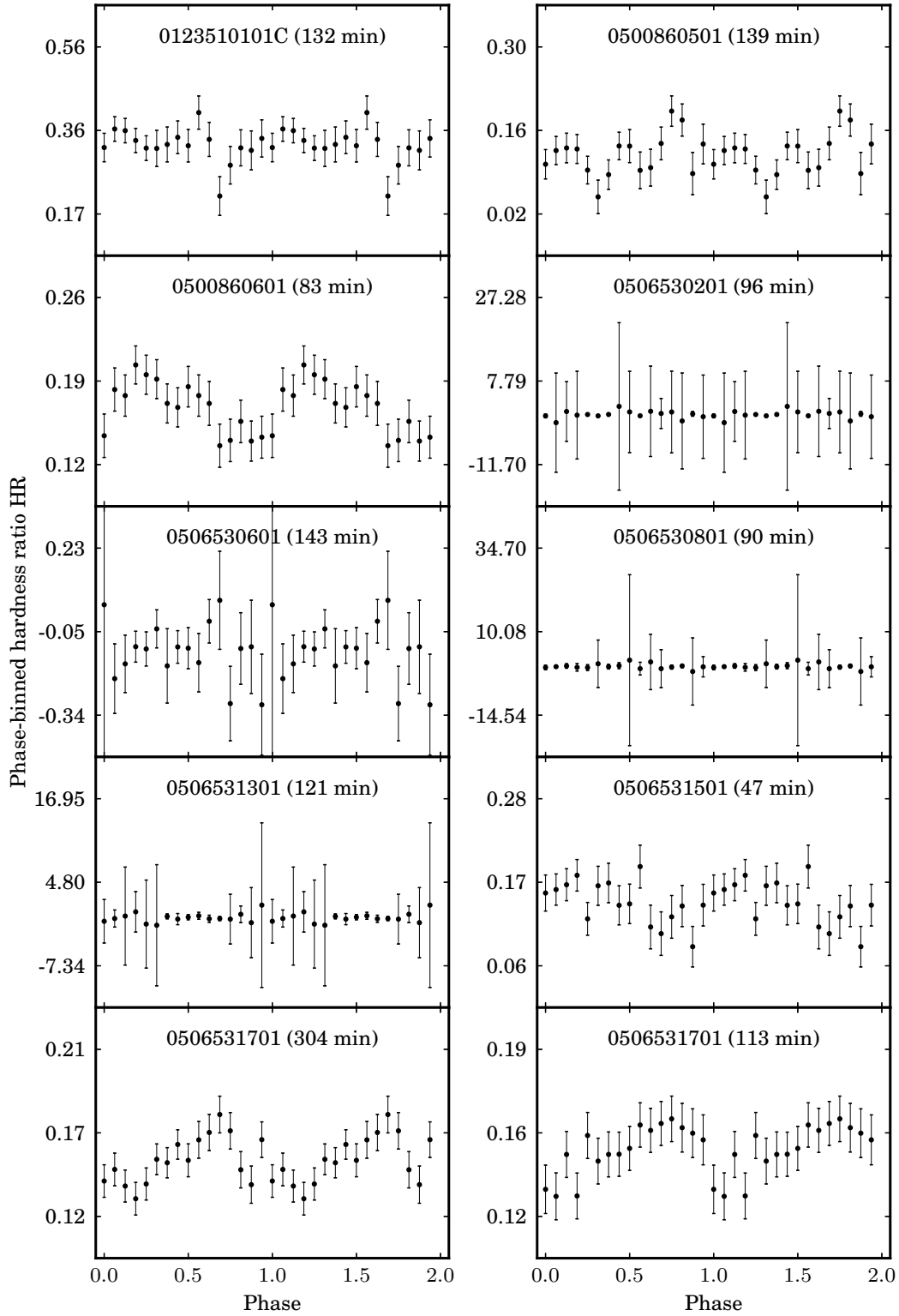


Fig. 4.23: The EPIC pn hardness ratio lightcurves folded on the longer periods listed in Table 4.8. The starting point for the phase folding was taken at the start of each dataset. Two cycles are plotted for clarity.

In Fig. 4.24, the following quantities are plotted together with the OGLE lightcurve: the mean EPIC pn count rate (CR) and hardness ratio (HR), as well as their Spearman correlation coefficient r_s for each observation, and also the mean value of the ~ 67 s and other detected X-ray periodicities. The epochs of the X-ray off-states are also indicated. Only the 2007-2009 observations are shown.

It is evident that the optical low states as illustrated by the OGLE lightcurve coincide with higher X-ray count rates, while the X-ray off-states and low states coincide with optical high states (see the discussion of the same data by Rajoelimanana et al., 2013). The X-ray and optical variability is not in the form of discontinuous jumps from one state to another, but smooth transitions can be observed, with the X-ray and optical flux invariably anti-correlated. X-ray “faint” states are observable when the optical magnitude is at intermediate values. The correlation between the CR and HR as determined earlier is also clearly visible in the graph.

In Fig. 4.25, the same OGLE lightcurve is provided, together with the optical AB magnitudes and colour indices derived from the OM data (see Table B.2). A zoomed-in version of the time from 20 March 2008 to 29 April 2008 which includes many observations, is provided in Fig. 4.26.

It can be seen in Fig. 4.25 that the modulations in the OGLE magnitudes are closely matched by the modulations in the OM AB magnitudes. This is especially evident in the optical rise and decline around $\text{BJD}_{\text{TDB}} \sim 2454570$, and the rise of ~ 0.6 mag around $\text{BJD}_{\text{TDB}} \sim 2454700$. The OM magnitudes in the different filters seem to follow the same trends in general, although small changes in the colour indices (i.e. changes in the shape of the near-UV continuum) do occur (see also Fig. 4.26).

4.7 Discussion

4.7.1 The nature of the X-ray and optical lightcurves

The long-term anti-correlation between the X-ray and optical flux in CAL 83 has been discussed by e.g. Greiner and Di Stefano (2002) (see §3.1.1). The onset of an X-ray off-state is explained by an increase in the photospheric radius of the WD at approximately constant luminosity, shifting the emission peak towards the far-ultraviolet, and thus increasing the optical and UV flux. This would also enhance the irradiation of the accretion disc, increasing the amount of reprocessed radiation from the disc and leading to an even higher optical/UV flux. Once the WD photosphere contracts again, the source will be detectable again in X-rays.

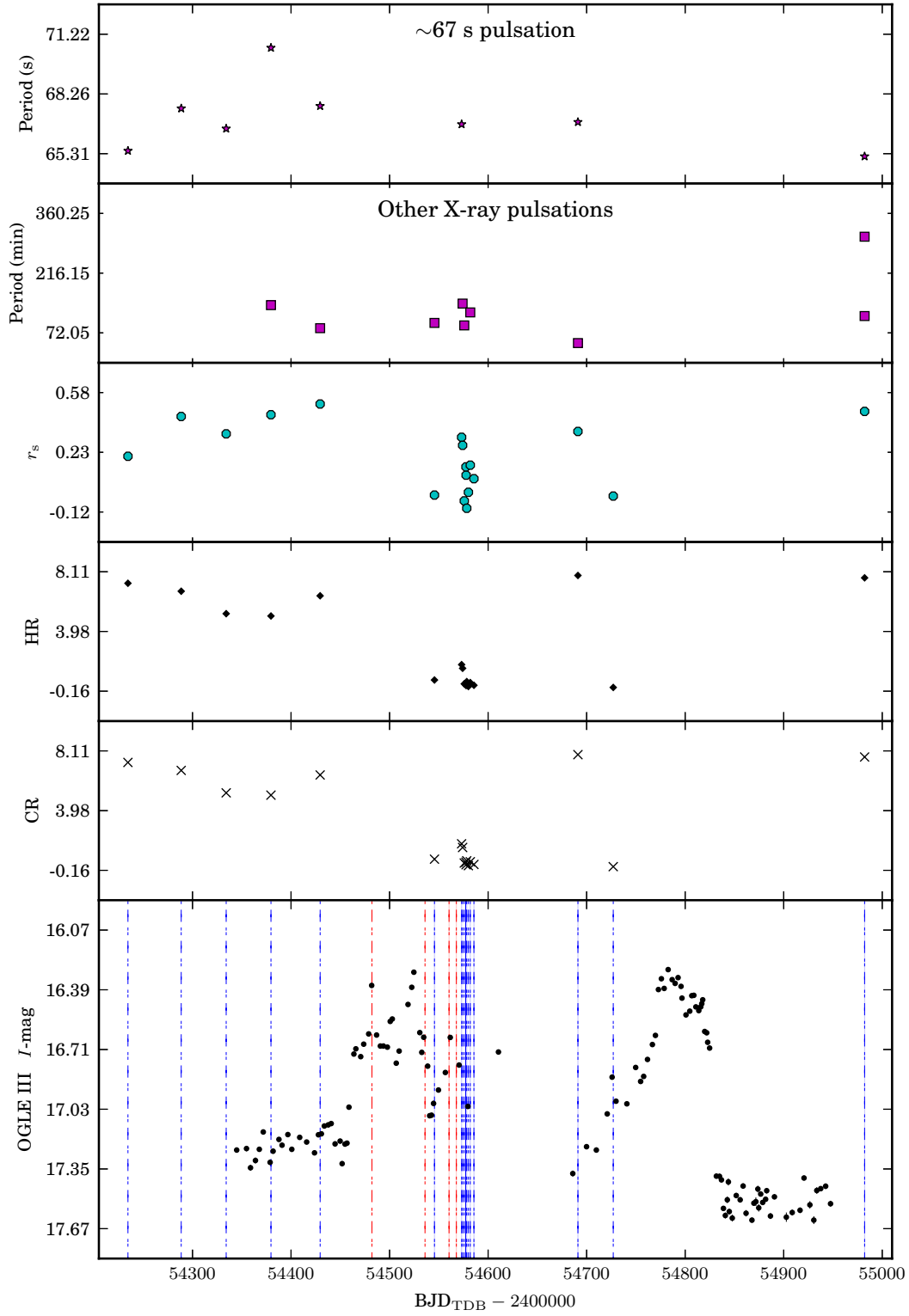


Fig. 4.24: The X-ray properties of CAL 83 derived from the 2007-2009 X-ray observations, together with the OGLE-III lightcurve. The times of the *XMM-Newton* observations are indicated on the OGLE plot by dashed lines: blue for the X-ray on-states, and red for the X-ray off-states. Error bars are included for all measurements except r_s , but are mostly smaller than the data points.

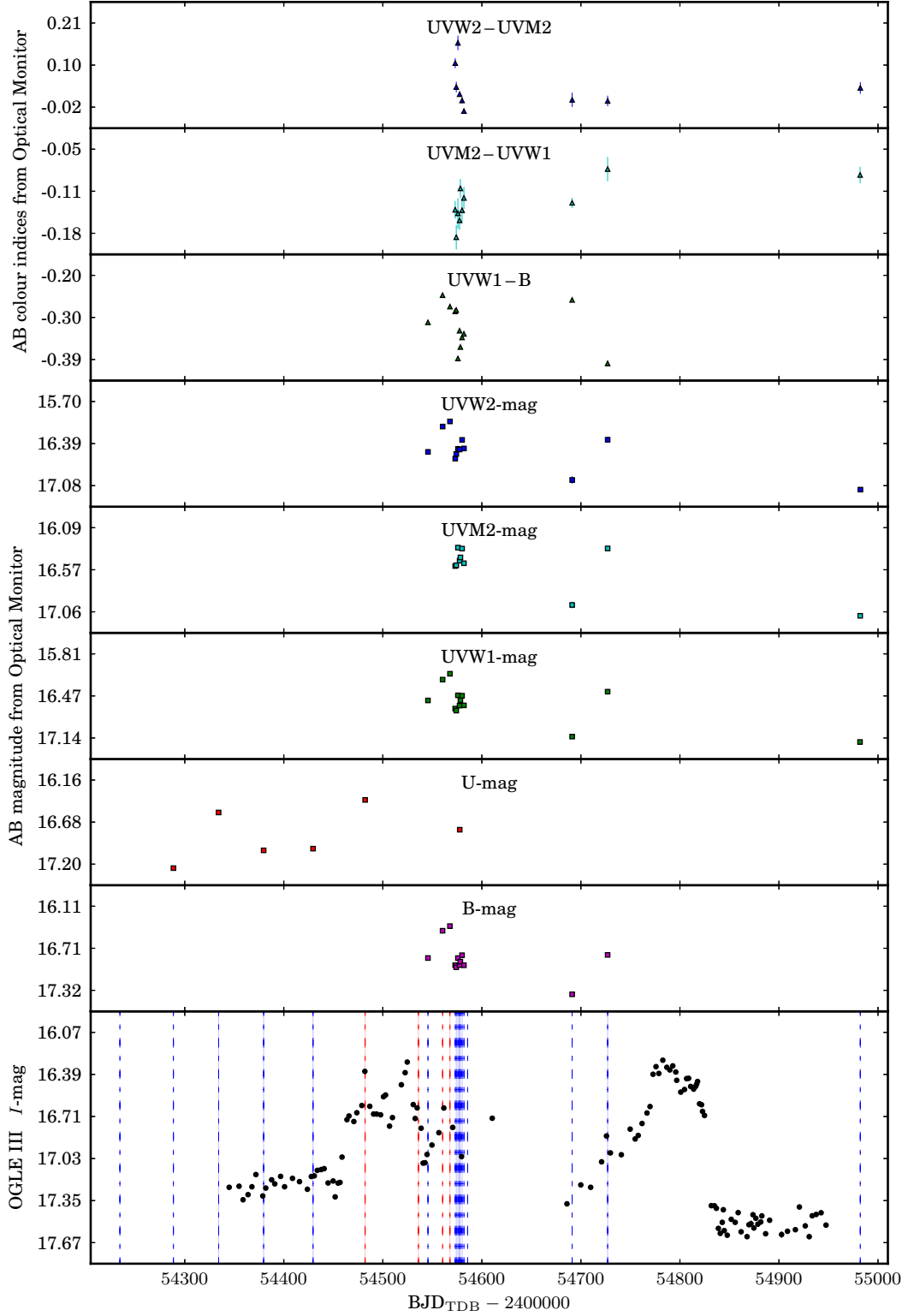


Fig. 4.25: The optical lightcurves of CAL 83 from OGLE and the Optical Monitor. The times of the *XMM-Newton* observations are indicated on the OGLE plot by dashed lines: blue for the X-ray on-states, and red for the X-ray off-states. Error bars are included for all measurements except r_s , but are mostly smaller than the data points.

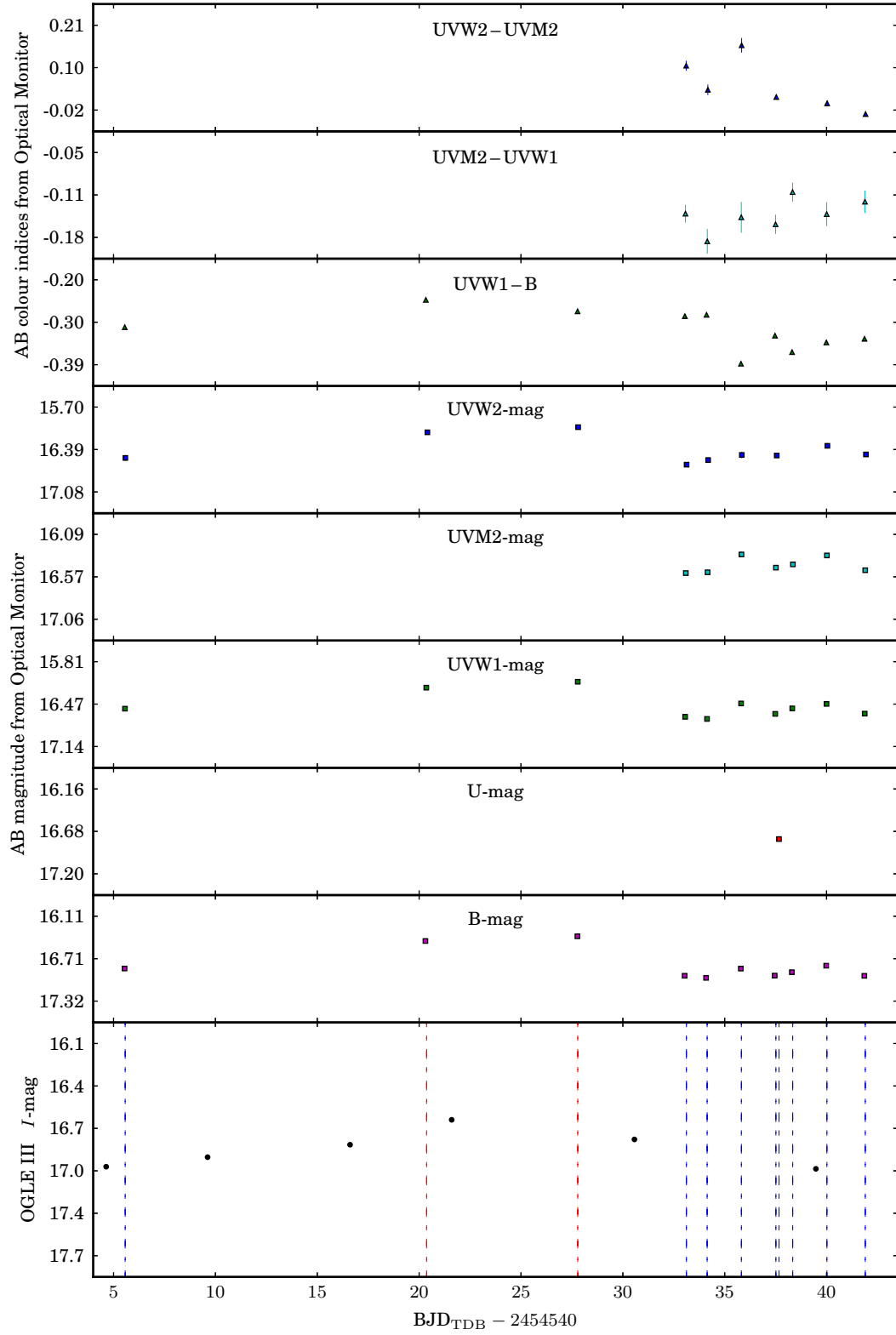


Fig. 4.26: The same as Fig. 4.25, but showing only the observations from 20 March 2008 to 29 April 2008.

The source RXJ0513 exhibits a similar anti-correlation, but on shorter time-scales. This has also been explained as an expansion/contraction of the WD photosphere by e.g. Reinsch et al. (2000), where a limit cycle involving changing disc viscosity due to changing amounts of irradiation sustains the quasi-regular cycles (see §3.2.1). Hachisu and Kato (2003a) proposed a slightly different limit cycle model for RXJ0513, where the onset of the X-ray low, optical high state is brought about by the accretion rate rising above the critical value $\dot{m}_{\text{crit}}$ defined in §2.6.3, causing an expanded WD photosphere as well as an accretion wind that obscures the supersoft X-rays, stripping off the outer layers of the donor and attenuating the mass transfer. The lowered accretion rate then enables the WD photosphere to shrink and the wind to cease, introducing an X-ray high, optical low state again.

CAL 83 has been observed extensively with *XMM-Newton*, with a total of 23 observations: one in April 2000 and the rest during the 2007 to 2009 time frame. Four of these observations were during an X-ray off-state, with 19 on-state observations exhibiting a large spread in X-ray count rates. The first observation, 0123510101, was performed in 3 segments, which were treated separately, yielding a total of 21 on-state datasets. These datasets have been analysed in detail in this chapter, focussing on the variability of the source.

Spectral fitting did not form part of the current investigation, but a hardness ratio was defined as a relative measure of the temperature associated with the X-ray spectrum. A highly significant correlation was found between the X-ray count rate and hardness ratio from observation to observation, i.e. on time-scales of days to years. This means that the effective temperature is higher when the source is brighter in X-rays, which is in good agreement with the long-term anti-correlation models.

The simultaneous near-UV and optical measurements obtained with the Optical Monitor (OM) on board *XMM-Newton* is also significantly higher during the X-ray off-states and low states, serving as further support for the photospheric expansion models. The correlated behaviour of the optical and near-UV magnitudes is in agreement with the similar findings of Bianchi and Pakull (1988) (see §3.1.2). The spectral energy distribution in the near-UV to optical region as inferred from the OM data (Fig. 4.7) indicates a very blue continuum in this waveband. A comparison of the OM flux values with the higher flux at even shorter wavelengths measured with the *Hubble Space Telescope* (Fig. 3.6) shows that the spectrum has a negative slope in the $\sim 1150\text{--}4500$ Å regime. The emission in the optical to near-UV waveband therefore represents the low energy tail of an emission component peaked at higher energies.

The individual X-ray observations exhibit significant variability in the count rate on

various time-scales from minutes to hours, with the standard deviations in the count rate varying from ~ 5 to 105% of the mean count rate. An investigation into the count rates and hardness ratios within the individual observations revealed a significant correlation between X-ray count rate and temperature even on short time-scales of minutes to hours. The periodograms derived from the hardness ratio variations did not exhibit significant peaks corresponding to either the ~ 67 s oscillation or the >500 s oscillations revealed in the count rate periodograms, with the exception of the strongest peak in the latter group. However, folding the hardness ratios on the periodicities derived from the count rates revealed weak quasi-sinusoidal modulations of the hardness ratio with the same periodicities and phasing as the count rate lightcurves in several of the observations, especially for the longer periods.

The OM instrumental magnitudes obtained from multiple exposures within a single observation exhibit small but significant variability, without any obvious relation to the X-ray count rate or temperature. The amplitudes of these variations are comparable to the ~ 0.1 mag of the orbital modulation. All the observations were significantly shorter than the ~ 1 d orbital period, and although the underlying orbital modulation is expected to be present, variability on shorter time-scales is also visible. Such modulations can represent any change in the near-UV emission regions in the system, perhaps changes in the structure or irradiation of the accretion disc rim.

4.7.2 A short white dwarf spin period due to disc torques?

The presence of a ~ 67 s X-ray periodicity in CAL 83 in X-ray observations spanning several years poses the possibility that it might be associated with the spin period of the accreting WD. It has been observed in individual runs and verified at a significance level of $\geq 3.5\sigma$ by an incoherent combination (stacking) of the periodograms of all the on-state observations. This short period is reminiscent of the short spin periods in cataclysmic variables like AE Aqr (33.08 s), V842 Cen (57 s) and V533 Her (63.633 s) and DQ Her (142 s). Longer spin periods have also been discovered in other supersoft X-ray sources (refer back to the discussion and references in §2.6.1).

The short rotation period in AE Aqr has been explained in terms of a high mass accretion history where the WD was spun-up by accretion disc torques during intervals of high mass transfer rate, during which it might then appear as a supersoft X-ray source as a result of stable nuclear burning in a shell on the WD surface (Meintjes, 2002; Schenker et al., 2002). The presence of supersoft X-ray emission in CAL 83 is also due to nuclear shell burning, placing CAL 83 in a similar situation that AE Aqr could have been in approximately ~ 10 Myr ago. Observational results indicate that the WD in CAL 83 has a mass of approximately $1.3 M_{\odot}$ Lanz et al.

(2005), indicating a long mass accretion history during which at least some of the accreted material was retained, even though a certain fraction of it may have been driven off in winds.

There are also other indications that CAL 83 may already have been in the super-soft phase for a considerable amount of time. The high ratio of C, N and O relative to He and H emission in the optical spectra of CAL 83 (see §3.1.3) may indicate a secondary star that has already passed through an extended period of mass transfer onto the WD. This notion is supported by the calculations of Chiang and Rappaport (1996), who considered the ionization evolution of the nebular region around CAL 83, and estimated that the source must have been radiating at a time-averaged luminosity of $\gtrsim 3 \times 10^{37}$ erg s $^{-1}$ for at least the past $\sim 10^5$ years. The high rate of mass accretion through an accretion disc over such a long time may therefore have spun up the WD in CAL 83 to a short rotation period of ~ 67 s (see the discussion in §2.6.1).

To investigate this possibility, one would need to take a closer look at some of the source parameters. According to the X-ray fits of Lanz et al. (2005), the luminosity of the source is $L_{\text{bol}} \sim 3.4 \times 10^{37}$ erg s $^{-1}$, and most of the other fits in Table 3.1 also indicate a luminosity of this order. The NLTE WD model atmosphere fits of these authors indicate a WD mass of $\sim 1.3 M_{\odot}$. According to Eq. (2.25), this yields a WD radius of 4.0×10^8 cm. Adopting an accretion rate of the order of $10^{-7} M_{\odot}$ yr $^{-1}$ (see below), the accretion luminosity according to Eq. (2.24) is $\sim 3 \times 10^{36}$ erg s $^{-1}$. The accretion luminosity is thus approximately an order of magnitude smaller than the bolometric luminosity, with most of the energy derived from nuclear burning in a shell on the WD surface.

The accretion rate can be estimated in different ways. The steady burning regime for hydrogen burning for a $1.3 M_{\odot}$ WD is between the limits $\dot{m}_{\text{std}} \sim 3.3 \times 10^{-7} M_{\odot}$ yr $^{-1}$ and $\dot{m}_{\text{cr}} \sim 6.8 \times 10^{-7} M_{\odot}$ yr $^{-1}$, according to Eq. (2.72) and (2.73) in §2.6.3, assuming a hydrogen fraction of $X = 0.7$. For hydrogen burning together with a small amount of helium burning, with the H and He fractions $X = 0.64$ and $Y = 0.35$ respectively, Eq. (2.75) yields an accretion rate of $\sim 10^{-7} M_{\odot}$ yr $^{-1}$ for a nuclear burning luminosity of 3×10^{37} erg s $^{-1}$.

If the optical/X-ray anti-correlation in CAL 83 is to be explained by a limit cycle model involving mass transfer at a rate comparable to the nuclear burning rate, the accretion rate may be close to $\dot{m}_{\text{cr}}$. This value may be considered as an upper limit to the actual accretion onto the WD, as a mass transfer rate in excess of this value will cause the excess material to be blown off in an accretion wind (Hachisu et al., 1996).

The optical spectra of CAL 83 exhibit variable broad Doppler shifted emission that has been ascribed to emission from an outflow originating in a precessing accretion disc. As discussed in §2.4.4, irradiation-induced disc warping can cause such precession in a disc. The accretion rate below which the disc is unstable to warping at $100R_{\text{wd}}$ is given by Eq. (2.50), and predicts an upper limit of $\sim 2 \times 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$ for a $1.3 \text{ M}_{\odot}$ WD with a nuclear burning luminosity of $\sim 3 \times 10^{37} \text{ erg s}^{-1}$. Therefore, within the limitations of the theory behind the equations utilized above, it appears as if irradiation-induced disc warping and an accretion rate close to the critical value can not occur simultaneously in CAL 83.

If the accretion rate is indeed $\sim 2 \times 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$, it would place CAL 83 below the steady burning strip, and in the regime of recurrent SSSs as described by Kahabka (1995). An X-ray off-state would then be caused by a decrease in the accretion rate, according to model (1) of Alcock et al. (1997) as discussed in §3.1.1. However, this would imply an X-ray low state that lags slightly behind the optical low state (associated with lower accretion rate), and not the anti-correlation between X-ray and optical that is observed. For the calculations below, an accretion rate of $\sim 10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}$ will now be adopted as an order of magnitude estimate.

If it is assumed that the WD had a rotation period as long as $\sim 1 \text{ d}$ at the beginning of the supersoft source phase (e.g. Kawaler, 2003), one can calculate the time required for the spin-up of the WD by accretion disc torques to a short spin period of 67 s by making use of Eq. (2.67). Assuming for the moment a WD magnetic field strength of $B_1 \sim 10^5 \text{ G}$ (see §4.7.4), the Alfvén radius can be obtained with Eq. (2.56):

$$R_{\text{M}} \sim 10^8 \left(\frac{B_1}{10^5 \text{ G}} \right)^{4/7} \left(\frac{R_1}{4.0 \times 10^8 \text{ cm}} \right)^{12/7} \times \left(\frac{M_1}{1.3 \text{ M}_{\odot}} \right)^{-1/7} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}} \right)^{-2/7} \text{ cm} . \quad (4.20)$$

This is smaller than the estimated Hamada-Salpeter radius of a $1.3 \text{ M}_{\odot}$ WD (Eq. (2.25)), which is $R_{\text{wd}} \sim 4 \times 10^8 \text{ cm}$, which means that the inner edge of a possible accretion disc will not be truncated by the WD magnetic field. We will show that this may have significant implications regarding the explanation of the observed $\sim 67 \text{ s}$ X-ray periodicity. The moment of inertia of the WD can be approximated as

$$I \sim 4 \times 10^{50} \left(\frac{M_1}{1.3 \text{ M}_{\odot}} \right) \left(\frac{R_1}{4.0 \times 10^8 \text{ cm}} \right)^2 \text{ g cm}^2 . \quad (4.21)$$

From Eq. (2.67), the resulting spin-up rate is

$$\dot{\Omega}_* \sim 2 \times 10^{-15} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}} \right) \left(\frac{M_1}{1.3 \text{ M}_{\odot}} \right)^{1/2} \times \left(\frac{R_M}{10^8 \text{ cm}} \right)^{1/2} \left(\frac{I}{4 \times 10^{50} \text{ g cm}^2} \right)^{-1} \text{ s}^{-2}. \quad (4.22)$$

The associated spin-up time-scale required to spin up the WD from a rotation period of $P_{*1} \sim 86400 \text{ s}$ to $P_{*2} \sim 67 \text{ s}$ is then

$$\tau_{\text{s-u}} = \frac{\Omega_{*2} - \Omega_{*1}}{\dot{\Omega}_*} = \frac{\frac{2\pi}{P_{*2}} - \frac{2\pi}{P_{*1}}}{\dot{\Omega}_*} = \frac{\frac{2\pi}{67 \text{ s}} - \frac{2\pi}{86400 \text{ s}}}{2 \times 10^{-15} \text{ s}^{-2}} \sim 10^7 \text{ yr}. \quad (4.23)$$

The mass transfer in supersoft sources typically takes place on the thermal time-scale of the donor, which is also $\gtrsim 10^7 \text{ yr}$ (see Eq. (2.21)). It is noted that a strong magnetic field of $B_1 \sim 10^7 \text{ G}$ yields $R_M \sim 2 \times 10^9 \text{ cm}$, with $\dot{\Omega}_* \sim 8 \times 10^{-15} \text{ s}^{-2}$ and $\tau_{\text{s-u}} \sim 4 \times 10^5 \text{ yr}$. A higher accretion rate would also bring about a higher spin-up rate. These estimates illustrate that the WD in CAL 83 could have been spun up by disc torques to a rotation period of $\sim 67 \text{ s}$ if it has been in the supersoft phase for a substantial amount of time. As discussed at the beginning of this subsection, other observational evidence does indeed indicate that this might be the case.

However, it has been shown that the $\sim 67 \text{ s}$ X-ray pulse in CAL 83 exhibits a drift of approximately 3 s around the mean of 67 s on time-scales of hours. Adopting source parameters as before, Eq. (2.68) shows that the time required to decrease a $\sim 67 \text{ s}$ spin period by 1 s is $\sim 21000 \text{ yr}$ and $\sim 5400 \text{ yr}$ for magnetic field strengths of 10^5 and 10^7 G respectively. It is therefore impossible for the rapid variations in the detected period to be due to accretion-induced spin-up and spin-down of the WD itself, due to its large moment of inertia.

Another obvious cause for modulations in an observed rotation period is the orbital motion of the WD (see §2.6.1). By assuming that the radial velocity curve of the He II $\lambda 4686$ line represents the WD motion, the radial velocity semi-amplitude has been estimated by several authors to be $K_1 \sim 30 \text{ km s}^{-1}$ (refer back to §3.1.3). With this value, Eq. (2.65) yields a maximum deviation from the mean of a 67 s spin period of 0.007 s . It is of course entirely possible that the He II $\lambda 4686$ modulation may underestimate the radial velocity modulation, but even increasing K_1 by two orders of magnitude still yields an expected period modulation of less than a second. Based on this alone, the variation in the $\sim 67 \text{ s}$ period can not be due to orbital motion.

Additionally, one can consider the longest *XMM-Newton* observation, 0506531701, which stretched over approximately one half of the orbital period (Fig. 4.16). If there was a direct correlation between the period value and the orbital motion, one would expect the period variation to take the shape of one half of a sinusoid. However, the pattern of the variability of the modulation in this observation does not suggest any correlation whatsoever with the orbital motion.

If the periodicity does represent the WD rotation period, it might be observed through a well developed envelope with an underlying nuclear burning shell at its base, surrounding the central degenerate core. The rotation of the envelope may not be quite synchronized (co-rotating) with the rapidly spinning WD, resulting in a slippage of the layers on the WD surface, and a spread in the observed ~ 67 s periodicity. In such a scenario, the actual WD rotation period may be somewhat shorter than 67 s. The actual detection of the periodicity then implies asymmetry in the brightness distribution, possibly because of the accretion being focussed in accretion arcs/curtains (see also §4.7.4).

The possibility was also investigated whether the ~ 67 s periodicity could represent a beat period between an even shorter rotation period and the Keplerian period of dense X-ray emitting gas close to the WD orbiting in the same direction as the WD, acting primarily as reflectors for an illuminating searchlight from the accreting WD. If the ~ 67 s oscillation is a beat between the real spin period and hot reflecting blobs orbiting next to the accreting WD, one can readily show (e.g. Patterson, 1979) that the spin period would then be

$$\frac{1}{P_*} = \frac{1}{P_b} + \frac{1}{P_K}, \quad (4.24)$$

where P_b represents the ~ 67 s beat period and P_K the Keplerian period of orbiting blobs close to the surface of the WD. Blobs orbiting the WD in the disc at distances R_K of the order $R_{WD} \leq R_K \leq 10^9$ cm will have Keplerian periods in the $P_K \sim 4\text{--}15$ s range, requiring WD rotation periods of $P_* \sim 4\text{--}12$ s (frequencies $\sim 80\text{--}270$ mHz) to produce the ~ 67 s beat period, where the lower limit represents the break-up period of a $1.3 M_\odot$ WD, $P_{b-u} = 3.8$ s (refer to Eq. (2.51) and Eq. (2.32)). However, extending the Lomb-Scargle analysis to higher frequencies by using bin sizes of down to 1 s revealed no consistent oscillation between 80 and 270 mHz.

4.7.3 Non-radial white dwarf pulsations?

Another possible, and perhaps very likely, explanation of the ~ 67 s periodicity is that it may represent non-radial g -mode pulsations in the WD (see the discussion in

§2.6.2). The presence of nuclear burning on the WD surface opens up the possibility of such pulsations driven by the ϵ -mechanism, in addition to the traditional $\kappa - \gamma$ mechanism.

Schmidtke and Cowley (2006) reported the discovery of a 38.4 min periodicity in the *Chandra* lightcurve of CAL 83, which they also ascribed to non-radial WD pulsations. It can be seen from Fig. 2 of their paper that the significance of the peak is only slightly above the 2σ significance level, quoted as $>97\%$. Among the more recent *XMM-Newton* observations analysed in this chapter, only 0506531501 had a significant peak near to this period, at 47.4 ± 11 min. However, there is also a weak peak in observation 0500860501 at this position which remained invariant under the different Lomb-Scargle analyses described in §4.5, even though its significance is lower than 99.73% in the original periodograms. Therefore, if the 38 min pulsation is real, then it does not seem to appear consistently. Modulations with periods ranging from ~ 90 to ~ 304 min were also detected in the more recent *XMM-Newton* data. These may also be related to non-radial WD pulsation modes.

The environment on the surface of the WD in an SSS is notably different than on the surface of a WD without nuclear burning or even accretion. Modulations in the accretion rate, nuclear burning rate and envelope size create a complex environment, and detailed modelling of the interactions taking place will help us to understand which pulsations will be triggered in such a scenario, and what type of variability can be expected in the observed periods. One must also bear in mind that non-radial oscillations are expected to exhibit multiple period peaks.

4.7.4 A LIMA-type model for the 67 s periodicity

It is interesting to compare the properties of the ~ 67 s period in CAL 83 with those of dwarf nova oscillations (DNOs). DNOs represent a type of quasi-periodic modulation seen in dwarf novae during outburst, and in other cataclysmic variables CVs with high accretion rates, but not in intermediate polars. DNOs typically have periods in the ~ 5 -40 s range, with variation amplitudes of $<1\%$ in the optical waveband. However, they have also been observed in X-rays, with much larger modulation amplitudes reaching up to 100% of the X-ray flux. Typical values of Q for DNOs are in the 10^3 to 10^7 range. DNOs also exhibit a period-luminosity relation, i.e. the period is slightly shorter at higher luminosity, associated with a higher accretion rate.

The ~ 67 s periodicity has a modulation semi-amplitude in the 2.5-8.6% range. With the subdivision and binning applied to the *XMM-Newton* lightcurves, Q -values from a few thousand down to ~ 60 was obtained for this periodicity. Therefore, the modulation period in CAL 83 is slightly longer, and the variability in the period somewhat

more rapid than is typically expected for DNOs. Although DNOs have so far only been sufficiently studied in WD systems without surface nuclear burning, considering the ~ 67 s periodicity in the context of a similar model as the one applied to DNOs may still be useful.

As explained in §2.7.4, the favoured interpretation of DNOs is that they may originate from a region close to the surface of a WD, where they represent the Keplerian period of an equatorial belt with a larger angular velocity than the WD itself. The details of this low-inertia magnetic accretor (LIMA) model to explain the DNO phenomenon has been developed by Warner (1995*b*) and Warner and Woudt (2002). Key elements of this model have been summarized in §2.7.4.

We now propose that the ~ 67 s periodicity in CAL 83 might originate in a similar equatorial region around the WD. The photospheric WD radius has been determined to be $(7.0 \pm 0.7) \times 10^8$ cm (Lanz et al., 2005), and this will be taken as an approximate outer limit for the equatorial belt. The Keplerian radius corresponding to a 67 s period is 2.7×10^9 cm (from Eq. (2.32)), while the Keplerian period at $\sim 7 \times 10^8$ cm is significantly lower, i.e. ~ 9 s. A belt with rotation period ~ 67 s at $\sim 7 \times 10^8$ cm thus rotates slower than the local Keplerian velocity, probably due to its coupling to a WD core with an even longer rotation period.

For the existence of such a belt with larger angular velocity than the WD, the magnetic field of the WD is perhaps too weak to rigidly couple this region to its core. An upper limit to the magnetic field strength can be calculated with Eq. (2.97). An estimate of the potential spin-up rate $\dot{\Omega}_*$ due to accretion torques first needs to be obtained. From Eq. (4.20) and (4.22), one can see that the value of $\dot{\Omega}_*$ is proportional to B_1 to the power $2/7$ (through its dependence on R_M), and therefore does not change too drastically with large changes in magnetic field. The value of $\dot{\Omega}_*$ changes by only one order of magnitude from 10^{-15} s^{-2} to 10^{-14} s^{-2} when varying the dipole field strength of the primary over the wide range from $B_1 \sim 7 \times 10^3$ G to $\sim 3 \times 10^7$ G.

Assuming $\rho_{\text{wd}} \sim 10^6 \text{ g cm}^{-3}$, the upper limit for the product of the interior magnetic field components to avoid rigid rotation according to Eq. (2.97), which was adopted from Katz (1975) (see the discussion in §2.7.4), is

$$B_r B_\phi \approx 6 \times 10^9 \left(\frac{\rho_{\text{wd}}}{10^6 \text{ g cm}^{-3}} \right) \left(\frac{\dot{\Omega}_*}{10^{-15} \text{ s}^{-2}} \right) \left(\frac{R}{7 \times 10^8 \text{ cm}} \right) \text{ G}^2. \quad (4.25)$$

Approximating the two components as being equal, and adopting $\dot{\Omega}_* \sim 10^{-15} \text{ s}^{-2}$, one obtains $B_r \sim B_\phi \lesssim 8 \times 10^4$ G, while for $\dot{\Omega}_* \sim 10^{-14} \text{ s}^{-2}$, the upper limit is $B_r \sim B_\phi \lesssim 2 \times 10^5$ G. In turn, the inner field strengths can be considered as an up-

per limit to the strength of the surface field B_1 . As an order of magnitude estimate, the primary field strength should therefore obey the limit $B_1 \lesssim 10^5$ G to allow the existence of a non-corotating belt structure at $R \sim 7 \times 10^8$ cm.

We will now investigate the possibility that the modulations in the ~ 67 s period may arise from the spin-up and spin-down of the equatorial belt. The maximum spin-up rate $\dot{\Omega}_b$ of the belt can be estimated by considering the period variations as illustrated in Fig. 4.16. The rapid decrease in the ~ 67 s period from 5 to 5.8 hours elapsed since T_0 in observation 0506531701 will be assumed to be representative of the most rapid change in the period of the equatorial belt (P_b), and will be used to estimate the maximum value of $\dot{\Omega}_b$. The period difference between the two segments is $\Delta P = (68.4 - 64.5)$ s ~ 4 s, and the time elapsed between them is $\Delta t = 3000$ s. Thus, $\dot{P}_b = \Delta P / \Delta t = 1.3 \times 10^{-3}$, yielding

$$\dot{\Omega}_b \sim 1.8 \times 10^{-6} \left(\frac{\dot{P}_b}{1.3 \times 10^{-3}} \right) \left(\frac{P_b}{67 \text{ s}} \right)^{-2} \text{ s}^{-2}. \quad (4.26)$$

In Eq. (2.67), the factor $(GM_1 R_M)^{1/2}$ represents specific angular momentum in a Keplerian rotating mass element (see Eq. (2.34)). The specific angular momentum of the belt magnetically linked to the WD is estimated by $Rv_{b,\phi}$, where $v_{b,\phi} = 2\pi R/P_b$ is its azimuthal velocity. The spin-up rate $\dot{\Omega}_b$ of the belt can thus be expressed as

$$\dot{\Omega}_b = \dot{m}_{\text{acc}} \left(\frac{2\pi R^2}{P_b} \right) I_b^{-1}, \quad (4.27)$$

or, solving for the moment of inertia I_b of the belt, one obtains

$$\begin{aligned} I_b &= \frac{\dot{m}_{\text{acc}}}{\dot{\Omega}_b} \left(\frac{2\pi R^2}{P_b} \right) \\ &\sim 1.6 \times 10^{41} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_\odot \text{ yr}^{-1}} \right) \left(\frac{\dot{\Omega}_b}{1.8 \times 10^{-6} \text{ s}^{-2}} \right)^{-1} \times \\ &\quad \left(\frac{R}{7.0 \times 10^8 \text{ cm}} \right)^2 \left(\frac{P_b}{67 \text{ s}} \right)^{-1} \text{ g cm}^2, \end{aligned} \quad (4.28)$$

which is almost 10 orders of magnitude smaller than that of the WD. This represents a maximum value for the inertia if the observed $\dot{\Omega}_b$ is to be explained by the spin-up of the belt. One can now proceed to estimate the mass of the belt. It will be assumed that the geometry of the belt can be described as an annular cylinder around the WD, with central axis perpendicular to the orbital plane. The outer radius R_{b2} of the belt will be taken as approximately equal to the photospheric radius of the envelope, i.e. $\sim 7 \times 10^8$ cm. The inner radius R_{b1} of the cylinder has a lower limit of

$R_{\text{wd}} \sim 4 \times 10^8$ cm (from the Hamada-Salpeter relation), although R_{b1} is probably larger than this. The moment of inertia is then given by (e.g. Halliday, Resnick and Walker, 2005, p. 253)

$$I_{\text{b}} = \frac{1}{2} m_{\text{b}} (R_{\text{b1}}^2 + R_{\text{b2}}^2) , \quad (4.29)$$

where m_{b} is the mass of the belt, which can be solved for as

$$m_{\text{b}} = \frac{2I_{\text{b}}}{R_{\text{b1}}^2 + R_{\text{b2}}^2} . \quad (4.30)$$

Adopting the radii described above yields $m_{\text{b}} = 2.5 \times 10^{-10} M_{\odot}$. According to Eq. (2.71) in §2.6.3, the critical envelope mass required for the ignition of nuclear burning is $m_{\text{env}}^{\text{crit}} \sim 9 \times 10^{-7} M_{\odot}$ for a $1.3 M_{\odot}$ WD accreting at $\sim 10^{-7} M_{\odot} \text{ yr}^{-1}$, which can be considered as a lower limit for the envelope mass for nuclear burning to occur. Therefore, if $m_{\text{b}} = 2.5 \times 10^{-10} M_{\odot}$ the equatorial belt represents only a small fraction of the total envelope mass.

Considering the extended, “fuzzy” nature of the WD envelope in SSSs, cloaking the nuclear burning shell, this scenario seems quite plausible. An equatorial belt at the envelope-disc boundary may have a rotational period of ~ 67 s as described above, while accreted layers closer to the WD may be rotating with slightly longer periods, yielding differential rotation between these weakly coupled layers at different radial distances. In such a case it is more likely that the inner radius of the belt is significantly larger than the WD radius. Increasing R_{b1} up to almost $R_{\text{b2}} = 7 \times 10^8$ cm (implying a thin shell at this radius) yields $m_{\text{b}} = 1.6 \times 10^{-10} M_{\odot}$, i.e. the presumed thickness of the shell does not change the resulting mass significantly.

The upper limit to the magnetic field strength to prevent rigid corotation of the belt with the WD was calculated earlier. However, the magnetic field needs to be strong enough to have a significant influence on the rotational motion of the material in the belt. In other words, although the coupling to the slower rotating WD core should not be too rigid, it should be strong enough to keep the belt from accelerating to the Keplerian period of ~ 9 s determined above. This condition may be evaluated by comparing the magnetic pressure to the ram pressure of the gas in the belt. A lower limit to the mass density of the belt can be obtained by evaluating the mass density of the inner disc. For an α -disc, the mass density is given by (e.g. Frank

et al., 2002, p. 93)

$$\rho_{\text{disc}} = 2 \times 10^{-4} \alpha^{-7/10} f^{11/5} \left(\frac{\dot{m}_{\text{acc}}}{10^{-7} \text{ M}_{\odot} \text{ yr}^{-1}} \right)^{11/20} \times \left(\frac{M_1}{1.3 \text{ M}_{\odot}} \right)^{5/8} \left(\frac{R}{7 \times 10^8 \text{ cm}} \right)^{-15/8} \text{ g cm}^{-3}, \quad (4.31)$$

where

$$f = \left[1 - \left(\frac{R_1}{R} \right)^{1/2} \right]^{1/4}. \quad (4.32)$$

Adopting a wide range of α -values ranging from 0.01 to 0.4 (e.g. Balbus and Hawley, 1998; King et al., 2007) yields values between 2×10^{-4} and $2 \times 10^{-3} \text{ g cm}^{-3}$ for the inner disc. Adopting the larger of these values as the density ρ_b of the belt, and calculating its azimuthal velocity $v_{b,\phi} = 2\pi(7 \times 10^8 \text{ cm})/67 \text{ s}$, the ram pressure can be approximated as $\rho_b v_{b,\phi}^2 \sim 9 \times 10^{12} \text{ erg cm}^{-3}$. The magnetic pressure associated with a $B_1 \sim 10^5 \text{ G}$ field is $B_1^2/8\pi \sim 4 \times 10^8 \text{ erg cm}^{-3}$, while the magnetic field required for the magnetic pressure to equal the ram pressure is $\sim 2 \times 10^7 \text{ G}$.

The ram pressure dominated flow will result in a significant winding up of the WD magnetic field frozen into the belt plasma. This will result in the generation of a significant toroidal external field. Assuming that the rotation period of the WD is substantially longer than that of the belt, the belt will carry the footpoints of the field lines around its orbit, stretching the field lines and creating a substantial toroidal field component $B_{b,\phi}$ in the vicinity of the belt (e.g. Aly and Kuijpers, 1990). It was shown by Warner and Woudt (2002) that the toroidal field generated by the belt is given by

$$B_{b,\phi} \sim 2 \times 10^5 \left(\frac{B_r}{10^5 \text{ G}} \right)^{1/2} \left(\frac{P_b}{67 \text{ s}} \right)^{-1/2} \left(\frac{m_b}{2 \times 10^{-10} \text{ M}_{\odot}} \right)^{1/4} \times \left(\frac{T}{5.5 \times 10^5 \text{ K}} \right)^{1/4} \left(\frac{x}{0.1} \right)^{-1/4} \left(\frac{y}{0.01} \right)^{-1/4} \left(\frac{M_1}{1.3 \text{ M}_{\odot}} \right)^{-1/4} \text{ G}. \quad (4.33)$$

The height of the belt (in the direction perpendicular to the orbital plane) is given by xR_1 . The quantity yg represents the effective gravity in the belt, where g is the surface gravity of the WD. The approximations $x \sim 0.1$ and $y \sim 0.01$ of Warner and Woudt (2002) were also used here. Adopting $T = 550000 \text{ K}$ according to the fits of Lanz et al. (2005), as well as $m_b \sim 2 \times 10^{-10} \text{ M}_{\odot}$, yields a toroidal field of $\sim 2 \times 10^5 \text{ G}$ for an original poloidal field of 10^5 G , which indicates a significant enhancement. It has been shown (Ghosh and Lamb, 1991, e.g. Meintjes and de Jager, 2000) that when the stretched toroidal field exceeds the poloidal WD field by up to 10 times, the field will become unstable to magnetic reconnection.

As the toroidal component $B_{b,\phi}$ of the magnetic field grows, reconnection of the increasingly stretched magnetic field lines through the tearing mode instability (see Priest and Forbes, 2000, Chapter 6) may cause the ejection of “bubbles” containing magnetic loops during the reconnection. These magnetic bubbles carry a significant amount of angular momentum, and are therefore centrifugally expelled towards outer regions of the belt, possibly expanding the effective outer radius of the belt, corresponding to a slightly longer observed period. Blobs of gas created via the Kelvin-Helmholtz instability may have a similar effect (e.g. Wang and Robertson, 1984), although the presence of substantial magnetic fields along the flow may attenuate these instabilities.

In dwarf novae, the increase in the DNO period at the end of the outburst has been explained as a rapid spin-down of the equatorial belt following a decrease in the accretion rate. Although SSSs are not subject to the viscosity-related hysteresis curve of dwarf nova outbursts, they probably undergo other processes that can modulate the mass transfer rate (as discussed in Chapter 3). However, these are not likely to occur on such short time-scales, and the spin-down of the belt is much more likely to be due to the reconnection process described above.

Alternatively, the rapid changes in the value of the ~ 67 s period within a single observation may be caused by the same mechanism proposed to explain the discontinuous jumps in the DNOs: the channelling of the accretion flow onto different accretion arcs in the belt (at different latitudes with different associated rotation periods) as a result of magnetic reconnection. In such a case, the disappearance and reappearance of the period on time-scales of hours (see Fig. 4.16) may be due to the accretion being channelled dominantly onto latitudes with the same period at some epochs (i.e. a well-defined period is observable), with accretion onto a wide range of latitudes with different periods at other epochs (i.e. a well-defined oscillation peak is not observable).

If spin-up and spin-down is not the primary cause of the period modulations, then the upper limit imposed on the inertia earlier in this section would not be so strict, and a larger fraction of the envelope mass could form part of the equatorial belt. However, it is probable that the variation in the ~ 67 s period might be caused by a combination of spin-up/spin-down and magnetic reconnection mechanisms.

Comparing the mean period and X-ray luminosity from observation to observation, marginal evidence has been found that the ~ 67 s period is on average somewhat shorter when the X-ray luminosity (and the corresponding temperature) is higher. This could simply be due to the fact that higher count rates and temperatures indi-

cate a slightly more contracted photospheric radius, with the belt at the envelope-disc boundary subsequently being at a slightly smaller radius with a smaller corresponding period. However, it is noted that the modulation in the ~ 67 s period from observation to observation over the years is not any larger than its modulation on time-scales of hours within a single observation.

Although the LIMA-type model proposed to explain the ~ 67 s period is still tentative, the results of the broad quantitative analysis above lead us to consider it as a very promising possibility. It definitely warrants further investigation and more detailed modelling of the possible existence of such an equatorial belt in SSSs.

Within the DNO framework, a class of longer period DNOs is also known (lpDNOs, see §2.7.4). Their periods are typically ~ 4 times longer than the “ordinary” DNOs, and the two types are often observed together. Also, it has been suggested by Warner (2004) that the lpDNO period has a direct relation to the spin period, with the spin period being ~ 2 times longer than the lpDNO.

If the ~ 67 s periodicity observed in CAL 83 is analogous to an “ordinary” DNO in a dwarf nova, one could hypothetically expect the inherent presence of additional oscillations close $P_{\text{lpDNO}} = 4 \times 67 \text{ s} = 268 \text{ s}$ (3.7 mHz) and $P_* = 8 \times 67 \text{ s} = 536 \text{ s}$ (1.9 mHz) in the lightcurves (although the presence of a DNO does not necessarily mean that the lpDNO and/or WD rotation period should also be detectable). The results of a Lomb-Scargle analysis in the 1-5 mHz region of the CAL 83 EPIC pn lightcurves with 10 s binning and detrended by subtracting a 2nd order polynomial fit, are shown in Fig. 4.27. A zoomed in version (in the 1-5 mHz range) of the average pn periodogram of the CAL 83 count rate in Fig. 4.12 is also shown in Fig. 4.28.

The only peak close to either 3.7 or 1.9 mHz that has a $>99.73\%$ significance level, is the one at ~ 1.7 mHz (580 s) in observation 0506530501. The region below 2 mHz has been studied in §4.5, and this ~ 1.7 mHz peak did not appear significant when using larger lightcurve bin sizes. However, it may be worthy of further investigation. None of the other observations contained an apparent peak at ~ 1.9 mHz. In observation 0123510101B and especially 0500860201 and 0506531201, a weak peak is visible approximately at 3.7 mHz, although below the $>99.73\%$ significance level. Although we can thus not confirm or exclude the presence of consistent periodicities at $P_{\text{lpDNO}} = 4 \times 67 \text{ s}$ and $2 \times P_{\text{lpDNO}}$, this is definitely something that could be investigated in more detail using new X-ray data.

The sources RXJ0513 and MR Vel have also been observed by *XMM-Newton*. Lomb-Scargle analyses of the X-ray lightcurves of these sources did not reveal any consistent short period modulations that may be associated with a highly spun up WD.

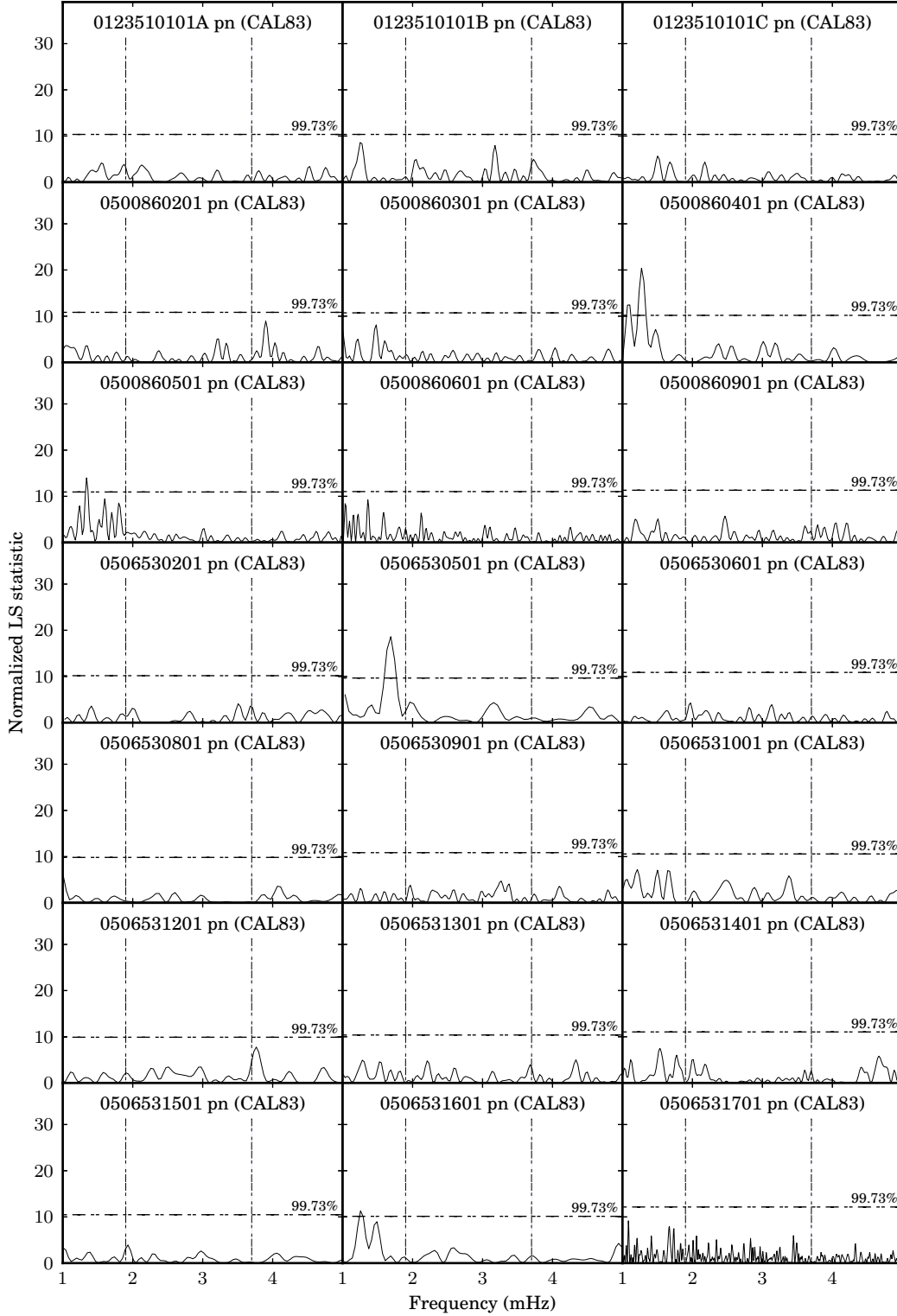


Fig. 4.27: Lomb-Scargle periodograms in the 1-5 mHz range of the EPIC pn data of CAL 83. The dashed vertical lines indicate the position of 1.9 and 3.7 mHz respectively.

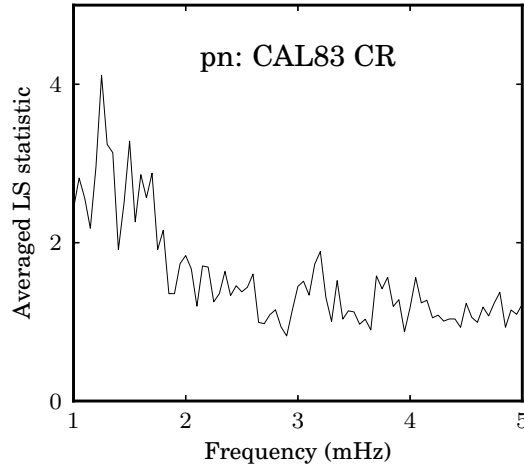


Fig. 4.28: The averaged Lomb-Scargle periodogram based on the EPIC pn broadband count rate, in the 1-5 mHz range. No prominent peaks at 1.9 and 3.7 mHz are apparent.

However, the possible presence of transient, low-frequency X-ray periodicities in RXJ0513, as observed in CAL 83, needs to be investigated further.

The observational evidence discussed in §3.1.3 and §3.1.2 indicates the presence of an accretion disc around the WD in CAL 83. As illustrated in this section, the presence and properties of such an accretion disc also play a key role in the interpretation of the X-ray modulations on time-scales from 67 s to days or years. Although the interaction of the inner disc with the WD envelope may be observed in soft X-rays, the regions further out in the disc emit in the optical waveband. In the next chapter, the properties of the accretion disc, as well as associated outflows, will be studied by making use of optical spectra obtained with the Southern African Large Telescope (SALT). The most extended spectroscopic study was undertaken for CAL 83, which is the main target of interest, but new spectra of the related sources RXJ0513 and MR Vel are also presented.

Chapter 5

Optical spectroscopy with SALT

Three close binary supersoft X-ray sources were observed with the Robert Stobie Spectrograph (RSS) on the Southern African Large Telescope (SALT): CAL 83 and RX J0513.9-6951 (hereafter RXJ0513) in the Large Magellanic Cloud, and the Galactic source MR Vel (also known as RX J0925.7-4758). All three of these sources are known to possess a wind or jet with velocities reaching several thousand kilometres per second. The main focus of this part of the study was to study the properties of the emission lines originating in the accretion disc and the outflows, i.e. the velocities, widths and intensities associated with these lines, and also the variability in these properties, with the purpose of characterizing the accretion process.

The chapter will be structured as follows: A discussion of the operation of the RSS will be presented in §5.1, and a summary of the observations in §5.2. In §5.3 the specific procedures used to reduce the RSS data are outlined. The spectroscopic calibration and analysis procedures are explained in §5.4 and §5.5 respectively. In the last three sections, the results and accompanying discussion of the sources, i.e. CAL 83, RXJ0513 and MR Vel, are presented separately. A general overview of the relevant properties of spectral lines is also provided for reference in Appendix C.

5.1 The Robert Stobie Spectrograph on SALT

The Robert Stobie Spectrograph (RSS) is the main work-horse instrument on the Southern African Large Telescope (SALT). Information about SALT can be found on the SALT websites for the public¹ and for astronomers², as well as in the paper of Buckley, Swart and Meiring (2006). The specifications of the RSS can also be found on these websites, in particular in the most recent call for proposals document (SALT Ast Ops, 2014), and also in the instrument papers of Burgh et al. (2003) and Kobulnicky et al. (2003). In this section, an overview of the telescope and the RSS

¹<http://www.salt.ac.za/>

²<http://astronomers.salt.ac.za/>

is provided, compiled from these sources.

A detailed discussion of the operation of astronomical spectrographs, as well as observing and calibration techniques in astronomical spectroscopy can be found in e.g. Howell (2006, pp. 135-166) and Massey and Hanson (2011).

5.1.1 The Southern African Large Telescope (SALT)

SALT is the largest single dish optical telescope in the southern hemisphere, and is located at the South African Astronomical Observatory (SAAO). The observatory is approximately 18 km east of Sutherland in the Northern Cape, with coordinates $20^{\circ} 48' 38.5''$ E; $32^{\circ} 22' 46''$ S, and an altitude of ~ 1800 m.

The primary mirror has a diameter of 11 m, and consists of 91 hexagonal segments, each of which is 1 m wide. The scientific payload is located at the primary focus. The primary mirror has a spherical shape, making the hexagonal segments interchangeable, but also introducing spherical aberrations into the images. The spherical aberration in the light beam reflected by the primary mirror is corrected by the Spherical Aberration Corrector (SAC), which is situated at the entrance of the payload.

SALT operates at a fixed altitude angle of 53° (observable range $47\text{--}59^{\circ}$) and can only rotate in the azimuthal direction to acquire targets. After acquisition, the tracker enables the tracking of the target by moving the optical payload over the focal plane of the mirror, without adjusting the azimuth angle. This places a limitation on the maximum single track time for a target, which is in the $\lesssim 1$ hour to $\gtrsim 3$ hour range, depending on the declination. It also implies a changing effective area during an observation, meaning that accurate absolute photometry and spectrophotometry are not SALT's recommended operational modes.

5.1.2 RSS operating modes

The Robert Stobie Spectrograph is mounted at the primary focus and provides low- to medium-resolution coverage of the 3200 to 9000 Å regime. Near-infrared capabilities of up to $1.7\text{ }\mu\text{m}$ are also planned for the future. It currently offers the following observing modes:

- Longslit spectroscopy
- Narrow-band imaging
- Multi-object spectroscopy
- Low-resolution Fabry-Perot imaging spectroscopy and tunable filter narrow-band imaging

- Single-etalon medium resolution Fabry-Perot imaging spectroscopy for studying emission lines
- High time-resolution spectroscopy

Medium- and high-resolution dual etalon modes are unavailable as at the time of writing, pending some rectifications. The RSS also has polarimetric capabilities, but these will only be commissioned after the problems with the Wollaston beam-splitter mosaic have been solved. The data discussed in this chapter were obtained in long-slit spectroscopy mode, which will be the focus of the discussion from here onwards.

5.1.3 Optical elements and layout of the RSS

For any telescope, the focal ratio or f/ratio is given by (e.g. Howell, 2006, p. 66)

$$f/ = \frac{\text{focal length of primary mirror}}{\text{primary mirror diameter}} . \quad (5.1)$$

A small f/ratio indicates that the parallel light pencil from a star is focussed “fast” by the telescope, in other words with a large positive vergence, while a larger f/ratio implies a “slower” focus, or a smaller positive vergence of the focussed pencil. An f/ratio can therefore also be used as an indication of the vergence of a light beam elsewhere in the optical system. The SALT SAC provides a fast f/4.2 beam at the focal plane, with a plate scale of 4.5 arcsec/mm, and a science field of view with an 8 arcmin diameter. An annulus extending to a diameter of 10 arcmin is used for guide stars.

The optical layout of the RSS is shown in Fig. 5.1. The focal plane is at position (1). A box adjacent to the focal plane stores a collection of slits, including 7 standard long-slits (see Table 5.1), 4 polarimetric long-slits, a few standard masks for imaging, polarimetry and calibration, as well as custom laser-cut masks for multi-object spectroscopy. The slit or mask appropriate for the operation mode is inserted into the focal plane during observations. In some imaging modes, no slit or mask is used.

Just behind the focal plane is a collimator system with several components, which yields a collimated beam with a diameter of 150 mm. The first is the field lens (2), which is 10 mm behind the focal plane, followed by a gap for the insertion of 1/2 or 1/4 waveplates for polarimetry at (3), or a glass blank for non-polarimetric modes. This is followed by 3 collimator lens groups at (4) (including a triplet with NaCl as the inner component!), a flat folding mirror (5) that changes the direction of the incident beam by 90°, and then the last doublet (6). For focussing, the first singlet and triplet can be moved together, and also the position of the second singlet

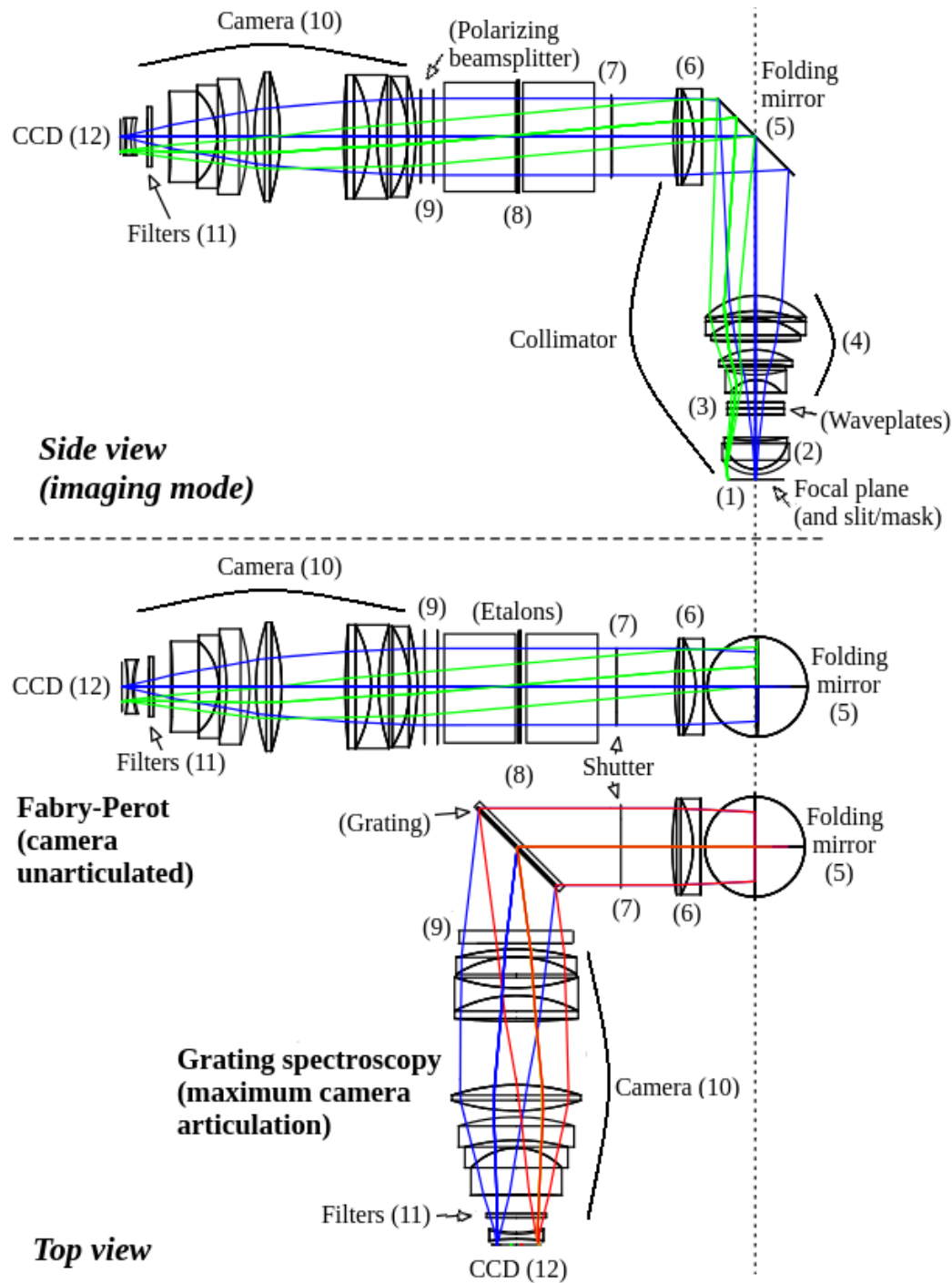


Fig. 5.1: The optical layout of the Robert Stobie Spectrograph. The top diagram shows a side view of the spectrograph in imaging mode, while the bottom diagram provides a top view of the layout for the following two cases: Fabry-Perot mode with the camera in the unarticulated position, and grating spectroscopy mode with the camera at maximum articulation. The grating angle and camera articulation can be varied according to the desired wavelength range. The blue and green rays in the imaging mode diagrams represent light pencils from an on-axis point and a 4 arcmin off-axis point respectively. The blue and red pencils in grating mode represent two different wavelengths to illustrate the dispersion. Parenthesized element names indicate that their presence or absence in the optical path depends on the observing mode. Numbers in parentheses are referred to in the text. (Adapted from Burgh et al., 2003, Fig. 1.)

Table 5.1: Available RSS slits for standard long-slit spectroscopy. All the slits have a height of 8 arcmin. (According to the SALT Principal Investigator Proposal Tool version 4.1.)

Longslit	Width (arcsec)
PL0060N001	0.6
PL0100N002	1
PL0125N001	1.25
PL0150N001	1.5
PL0200N001	2
PL0300N001	3
PL0400N001	4

Table 5.2: Available gratings on the RSS (Adopted from SALT Ast Ops, 2014, pp. 31-32.)

Grating	Wavelength coverage (Å)	Usable angles (°)	Bandpass per tilt (Å)	Resolving Power (for a 1.25" slit)
PG0300	3700-9000		3900/4400	250-600
PG0900	3200-9000	12-20	~ 3000	600-2000
PG1300	3900-9000	19-32	~ 2000	1000-3200
PG1800	4500-9000	28.5-50	1500-1000	2000-5500
PG2300	3800-7000	30.5-50	1000-800	2200-5500
PG3000	3200-5400	32-50	800-600	2200-5500

relative to the former group.

After the collimator and before the dispersers, there is a Prontor 150 mm shutter (7). In imaging mode, there will not be any dispersive element in the beam, while for Fabry-Perot imaging one or two of three available etalons is placed in the optical path at (8). For longslit spectroscopy, multi-object spectroscopy or spectropolarimetry, one of six available gratings will be mounted at (8) as the dispersive element. The disperser is followed by a beamsplitter for polarimetry mode (9).

The six available gratings and their properties are listed in Table 5.2. One is a regular transmission grating (pg0300), and the other five are volume-phase-holographic (VPH) gratings. Ordinary diffraction gratings diffract the light by means of their periodic surface structure, while a VPH grating diffracts the light by variations in refractive index within its volume (Barden et al., 2000). VPH gratings have many advantages, including a high diffractive efficiency, a reduction in the amount of scattered light and low induced polarization effects. The efficiencies of the RSS VPH gratings at different wavelengths and resolution values are shown in Fig. 5.2.

The angle between the incident beam and the normal to the grating surface is called the angle of incidence, also known as the *grating angle*. The efficiency of a VPH grating varies with the grating angle, making it possible to access a large wavelength range with a single grating by simply adjusting the grating angle. The grating is mounted on a rotatable stage, providing access to a large range of angles. Increasing the grating angle will also cause an increase in resolution.

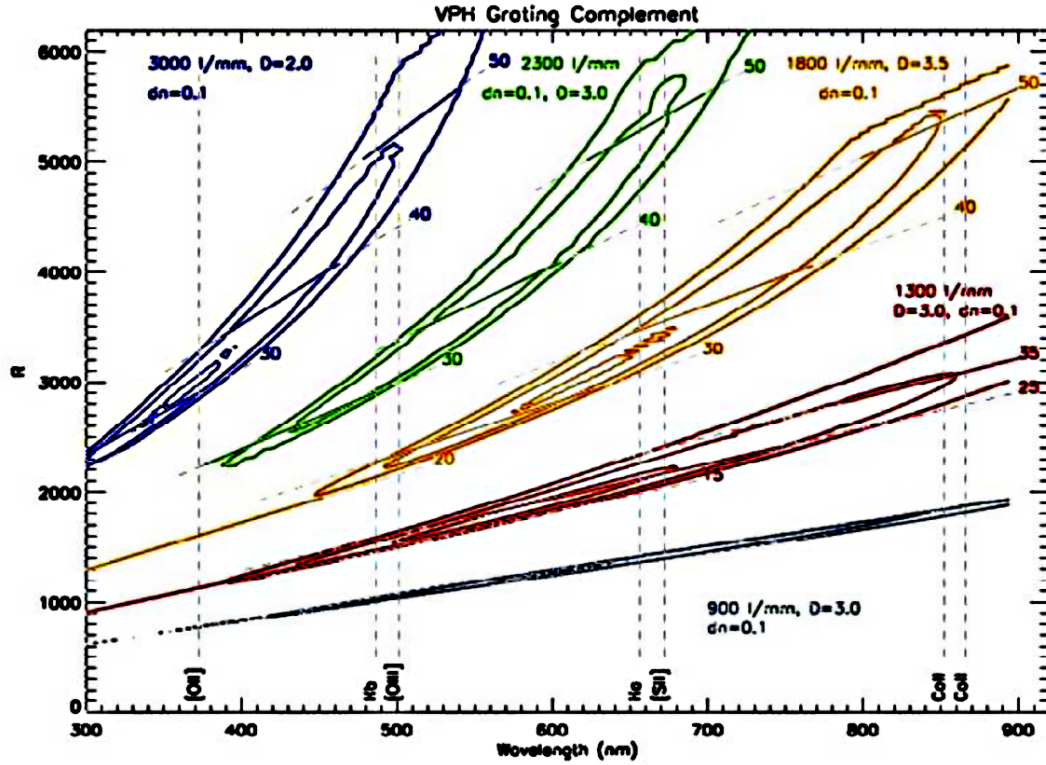


Fig. 5.2: RSS VPH grating efficiencies in terms of resolution (R) and wavelength. The contours indicate efficiencies of 90%, 70% and 50%. “D” indicates the grating thickness in μm and “dn” the modulation in the refractive index. Most of the VPH grating efficiencies peak at $> 90\%$, which is appreciably higher than the efficiencies of regular diffraction gratings. (Adopted from SALT Ast Ops, 2014, Fig. 4, p. 6.)

The camera is mounted in a tube that can articulate around an axis coinciding with the rotation axis of the grating. The camera position is described by the camera angle, which represents the angle between the beam incident on the grating, and the optical axis of the camera. For imaging and Fabry-Perot modes, the camera will be in the unarticulated position (the two top diagrams in Fig. 5.2), while the use of a grating will require camera articulation (like in the bottom diagram in Fig. 5.2). The VPH efficiency is usually the maximum at the Littrow condition, i.e. where the angles of incidence and refraction are equal, which requires the camera angle to be twice the size of the grating angle.

The camera (10) contains 9 lenses in 4 groups, and yields an $f/2.2$ beam and a plate scale of 8.6 arcsec/mm. The singlet and triplet are moved together to focus the camera. All the surfaces are spherical, except for the aspheric first surface. The last lens is a field flattener that also forms the CCD dewar window. The gratings are used only in first order, with second-order contamination prevented by using order-blocking filters at (11). Five order-blocking filters are available: one clear (PC00000), three UV blocking (PC03200, PC03400 and PC03850), and one blue

blocking (PC04600). There are also 40 available interference filters for Fabry-Perot observations and narrow-band imaging.

The detector is a 3×1 array of three E2V 44-82 CCD chips with 2048×4096 pixels each (2048 columns and 4096 rows). The pixel size is $15 \mu\text{m}$, yielding a plate scale of $0.13 \text{ arcsec/pixel}$. The separation between the chips is $225 \mu\text{m}$, causing two chip gaps that need to be taken into account when planning observations. The chips are cooled by a closed cycle Cryotiger, making the dark current insignificant. Readout takes place via 6 amplifiers: 2 for each CCD chip. Each amplifier reads out 1024 columns.

There are two gain settings: a “Bright” setting of $3.2 \text{ e}^-/\text{ADU}$ that provides a large dynamic range, and a “Faint” setting of $1.5 \text{ e}^-/\text{ADU}$ to use for fainter targets. Two readout speeds are available: “Slow” and “Fast”. In normal readout mode with a binning of 2×2 , the entire mosaic can be read out in 11.2 s in “Fast” mode, or in 17.7 s in “Slow” mode (according to the SALT Principal Investigator Proposal Tool version 4.1.). A readout noise of $5 \text{ e}^-/\text{pixel}$ is associated with the fast mode, with a lower readout noise of $\sim 3 \text{ e}^-/\text{pixel}$ for the slow mode. For observations requiring high time resolution, frame transfer or slot mode may be used, as well as the currently unavailable charge shuffling mode.

An illuminated screen, inserted below the SAC, is used to obtain flat fields. A selection of arc lamps are available for wavelength calibration (Kniazev, 2009)³. The arc lamps are situated in the prime focus payload of SALT, and the light is fed to the calibration system by liquid light guides.

5.2 Observations with the RSS

The optical spectra that will be presented in this chapter were obtained with the RSS in longslit spectroscopy mode on various nights during the 2011 to 2013 observing cycles. The log of these observations is presented in Table 5.3, and the relevant instrumental configurations in Table 5.4. As these targets are relatively faint (V -magnitude around 17), the “Faint” gain setting was used, together with the “Slow” readout speed to obtain a lower readout noise.

In the following sections, the reduction, calibration and analysis of these spectra are described. IRAF (the Image Reduction and Analysis Facility)⁴ version 2.15.1a was used, using the PyRAF environment version 2.1.6 with Python version 2.7.3. The spectral fitting described in §5.5 was performed with Fityk version 0.9.8, a curve

³<http://pysalt.salt.ac.za/lineatlas/>

⁴<http://iraf.noao.edu/>

Table 5.3: RSS longslit observations of CAL 83, RXJ0513 and MR Vel. See Table 5.4 for a description of the instrument configurations. The principal investigator for all the proposals was A. Odendaal, with P.J. Meintjes, P.A. Charles and A.F. Rajoelimanana as co-investigators.

Target	Proposal ID	Start date & time (UTC)		Exposures	Configuration
CAL 83	2011-3-RSA-017	2011-11-26	01:31:20	$6 \times 37 \text{ s} = 222 \text{ s}$	A
			01:55:47	$6 \times 37 \text{ s} = 222 \text{ s}$	B
	2012-1-RSA-019	2012-09-11	02:31:30	$2 \times 183 \text{ s} = 366 \text{ s}$	C
			02:48:33	$2 \times 183 \text{ s} = 366 \text{ s}$	D
	2012-2-RSA.UKSC-006	2012-10-29	00:25:11	$3 \times 183 \text{ s} = 549 \text{ s}$	A
			01:38:35	$3 \times 183 \text{ s} = 549 \text{ s}$	D
		2013-01-05	23:48:44	$2 \times 416 \text{ s} = 832 \text{ s}$	E
			19:09:09	$2 \times 432 \text{ s} = 864 \text{ s}$	E
			21:01:15	$2 \times 432 \text{ s} = 864 \text{ s}$	E
			18:39:13	$2 \times 432 \text{ s} = 864 \text{ s}$	F
	2013-1-RSA.UKSC-002	2013-03-21	19:08:01	$2 \times 320 \text{ s} = 640 \text{ s}$	F
			01:48:46	$2 \times 500 \text{ s} = 1000 \text{ s}$	E
			02:57:31	$2 \times 600 \text{ s} = 1200 \text{ s}$	E
			02:25:49	$2 \times 500 \text{ s} = 1000 \text{ s}$	E
RXJ0513	2012-1-RSA-019	2012-08-27	03:29:41	$2 \times 182 \text{ s} = 364 \text{ s}$	C
			03:40:47	$2 \times 182 \text{ s} = 364 \text{ s}$	G
		2012-10-29	02:07:47	$3 \times 182 \text{ s} = 546 \text{ s}$	A
			02:23:55	$3 \times 182 \text{ s} = 546 \text{ s}$	G
	2013-1-RSA.UKSC-002	2013-09-18	02:48:31	$2 \times 499 \text{ s} + 300 \text{ s} = 1298 \text{ s}$	E
			01:17:12	$3 \times 499 \text{ s} = 1497 \text{ s}$	E
			02:33:24	$2 \times 499 \text{ s} = 998 \text{ s}$	E
			2013-09-25	$2 \times 499 \text{ s} = 998 \text{ s}$	E
MR Vel	2012-2-RSA.UKSC-006	2013-03-21	23:07:42	$2 \times 423 \text{ s} = 846 \text{ s}$	E
			2013-04-29	$2 \times 423 \text{ s} = 846 \text{ s}$	E

Table 5.4: Configurations used for the RSS longslit observations of CAL 83, RXJ0513 and MR Vel listed in Table 5.3. The PG0900 grating and a 2×2 pixel binning was used for all observations. The gain was set to “Faint”, and the readout speed to “Slow”. The numbers in parentheses after the arc lamp specification in the last column indicates the average rms error applicable to the wavelength calibration for the observations utilizing that particular configuration, as well as the average number of arc lines used in the wavelength solution.

Configuration	Grating angle ($^{\circ}$)	Range ($\text{\AA}$)	Slit width (arcsec)	Resolution ($\text{\AA}$)	Filter	Arc correction
A	12.13	3040-6170	0.6	~ 2.3	PC00000	ThAr(0.79 $\text{\AA}$; ~ 96)
B	19.63	5875-8894	0.6	~ 2.2	PC04600	Xe (0.46 $\text{\AA}$; ~ 67)
C	12.13	3060-6180	0.6	~ 2.3	PC00000	HgAr(0.67 $\text{\AA}$; ~ 13)
D	19.63	5880-8900	0.6	~ 2.2	PC04600	Ne (0.21 $\text{\AA}$; ~ 53)
E	14.75	4050-7140	1.0	~ 3.8	PC03850	Ar (0.64 $\text{\AA}$; ~ 65)
F	12.13	3080-6180	1.0	~ 3.8	PC00000	Ar (0.63 $\text{\AA}$; ~ 55)
G	19.63	5880-8890	0.6	~ 2.2	PC04600	Ne (0.22 $\text{\AA}$; ~ 58)

fitting and analysis program⁵ (Wojdyr, 2010).

5.3 RSS data reduction

For a basic introduction to the properties of CCD cameras, as well as the necessary calibrations related to CCD data, the reader is referred to Appendix D. In this section, the focus will be on the reduction of the raw RSS data files. SALT data is distributed to the PI in two folders: one containing the raw data files, and one containing the product files obtained after running pipeline tasks in PySALT, which is the official PyRAF-based software package for SALT data reductions. The analysis reported here started with the raw data files, and tasks in the **saltred** package of PySALT version 0.47 were used to apply the pipeline processing forming the first five reduction steps (§5.3.2 to §5.3.6).

Information about PySALT were found on the SALT astronomers' website⁶, on the SALT Wiki⁷, and in the help files of the different PySALT tasks. Massey (1997) provides an excellent guide on CCD reductions with IRAF in general.

5.3.1 The FITS file format

FITS stands for Flexible Image Transport System, and refers to a type of file that consists of a number of segments called Header/Data Units (HDUs) (e.g. Pence, 2014⁸). The first HDU is the primary HDU, while the other HDUs are referred to as extensions. As the name indicates, each HDU can contain a header unit and a data unit. In the header unit (usually simply referred to as the header), a number of keywords are stored with important information related to the HDU and its processing history. The data unit contains an array or table with the numerical data, e.g. a spectrum, image or data cube.

Each RSS exposure is stored in a single FITS file consisting of 7 HDUs: the primary HDU has only a header containing a list of keywords relevant to the observation, while the other 6 ADUs contain the information associated with the 6 readout amplifiers. Each of the 6 extensions contains a data unit with the image read out from the associated amplifier, two per CCD, as well as a header with keywords specific to that amplifier.

⁵<http://fityk.nieto.pl>

⁶<http://astronomers.salt.ac.za/software/pysalt-documentation>

⁷<https://sciencewiki.salt.ac.za/index.php/PySALT>

⁸http://fits.gsfc.nasa.gov/fits_primer.html

5.3.2 Pre-processing

The **saltred** task **saltprepare** was used to prepare the data for processing. This task verified that the data file format and header keywords were compliant with IRAF standards, and added several necessary keywords that are not written to the files at the telescope.

5.3.3 Gain corrections

As discussed in §5.1.3, the gain of the RSS CCDs can be set to “Bright” or “Faint”. However, the six different amplifiers have slightly different actual gain factors, and the **saltred** task **saltgain** was used to perform a gain correction to all the CCD frames. For each separate amplifier in an exposure, the task read the appropriate gain value from the FITS extension header keyword, and multiplied the image associated with the amplifier by this factor. However, the gain factors also vary with time by 1-10%, making additional corrections necessary⁹ (see §5.3.8).

5.3.4 Cross-talk corrections

As mentioned before, each RSS CCD chip has two readout amplifiers. A small amount of cross-talk ($\sim 0.1\%$) occurs between the two amplifiers, and can be seen as “ghosts”, i.e. faint mirror images of bright sources over channel boundaries. This effect was corrected with the **saltred** task **saltxtalk**. This task subtracted a scaled image of one amplifier from the neighbouring image. The appropriate scale factors for the different amplifiers were read from the header keywords.

5.3.5 Bias subtraction

The FITS extension of each amplifier contains an image of 562×2051 pixels (after pre-binning of 2×2), including 50 overscan columns. The overscan pixels are in the regions [1:25,1:2051] and [538:562,1:2051] (in the format [column range, row range]). The data section is therefore [26:537,1:2051], and is indicated in the DATASEC header keyword. The default regions to use for determining the overscan are also stored in the headers under the keyword BIASSEC. The BIASSEC for amplifiers 1, 3 and 5 is [539:561,1:2051], and for 2, 4 and 6 it is [2:24,1:2051]. (Note that although the term “row” is used here, IRAF refers to rows as “lines”.)

The overscan correction was performed by using the **saltred** task **saltbias**. This task employed the smoothed overscan vector method described in §D.5. The different columns in the BIASSEC region were averaged, and a third-order polynomial was fitted to the resulting single column vector. The best fit function was subse-

⁹P. Vaisanen, private communication

quently subtracted from each column in the image.

The **saltbias** task also trimmed off the overscan regions from the images, yielding six 512×2051 extensions for each exposure. In cases where a master bias frame was available, **saltbias** was also used to subtract the master bias from the target, arc and flat frames.

5.3.6 Mosaicing

For a single exposure, the 6 images from the 6 amplifiers were stitched together by the **saltred** task **saltmosaic**. The task first joins the amplifiers from the same CCDs to obtain 3 sets of 2 joined images. After this, the 3 sets are joined together with the chip gaps between them, and during this process the outer sets are also rotated relative to the central one during a geometrical translation aimed at correcting distortion effects. The result is a FITS file with two HDUs: the primary HDU still containing only a header, and the extension containing a single two-dimensional 3170×2052 mosaiced image, as well as additional keywords.

As an example, the first CAL 83 spectrum obtained on 5 January 2013 is shown in Fig. 5.3. In this spectral image, the *dispersion direction* is in the horizontal or x -direction. In other words the dispersed beam of light from the diffraction grating is spread over the detector in such a manner that the blue is on the left, ranging to red on the right hand side. Therefore, each vertical pixel column with constant x on the image represents a different wavelength element. For this particular spectrum, wavelengths extending from 4050 to 7140 Å were dispersed over the 3170 columns. Referring back to the top view in grating spectroscopy mode at the bottom of Fig. 5.1, one can see the blue light being focussed on the left hand side of the CCD, and the red light on the right hand side. Therefore, the dispersion direction is parallel to the plane of the page, extending from small x -values on the left to large x -values on the right.

In a spectrograph, the orientation of the slit is perpendicular to the dispersion direction. With reference to Fig. 5.3, one can therefore imagine the slit in a vertical orientation earlier in the optical beam. The light from each element on the sky along the slit will be dispersed onto the CCD by the grating. Each horizontal row across the image therefore represents the dispersed spectrum of an element on the sky. For the data in Fig. 5.3, a binning of 2×2 was used, and with a plate scale of 0.13 arcsec/pixel, this means that each of the 2052 horizontal rows with constant y represents the spectrum of a 0.26 arcsec on-sky region on the slit. Consequently, the vertical direction is called the *spatial* or *cross-dispersion* direction. The spatial direction on the CCD in Fig. 5.1 (top view, grating spectroscopy mode) is therefore

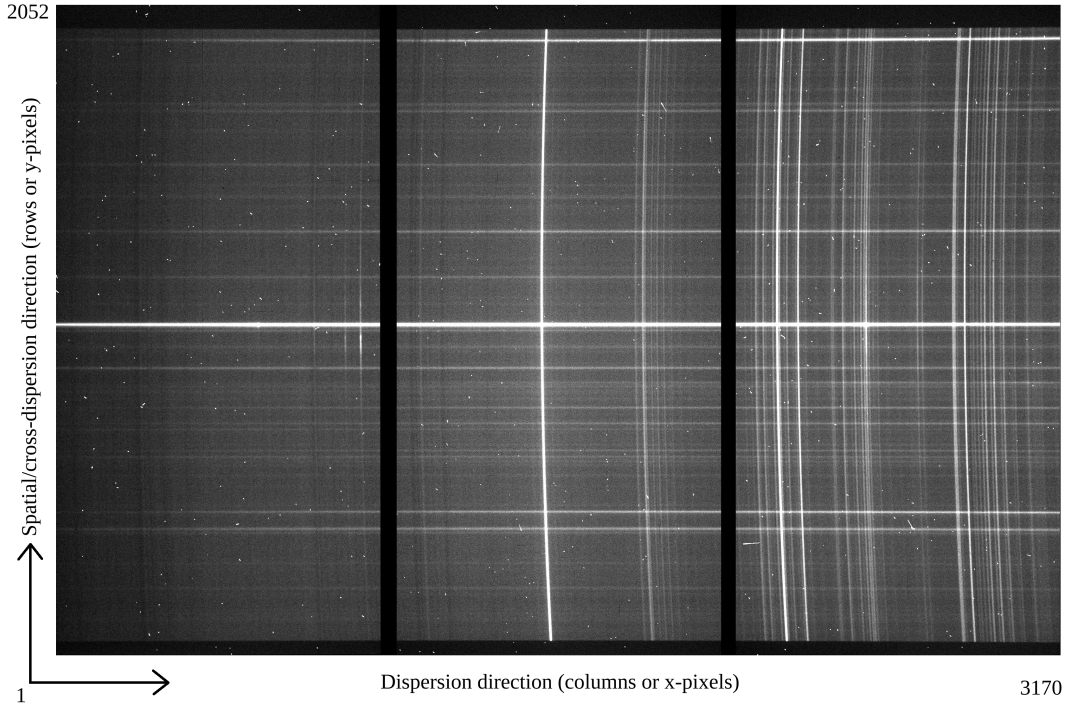


Fig. 5.3: Spectral CCD image of CAL 83 obtained with the RSS on 5 January 2013 (first exposure). The wavelengths increase from left to right in the dispersion direction. The vertical cross-dispersion direction represents the spatial range extending over the 8 arcmin field of view. The bright horizontal trace across the centre represents the target trace.

perpendicular to the page. The bright horizontal trace in Fig. 5.3 centred on row 1045 is the trace for CAL 83, and is spread over approximately 15 rows.

The bright vertical lines extending over the whole image represent line emission extending over the length of the slit. Most of these are atmospheric emission lines, but they can include any extended emission in the observed region. In this case, they will include background emission lines from the LMC. The random bright spots on the image are primarily caused by cosmic rays, but may also include a few CCD hot pixels.

In the discussion above, it was assumed that the dispersion direction is perfectly parallel to the x -direction, and the spatial direction perfectly parallel to the y -direction. It should be noted that distortions in the system can cause the dispersion and spatial directions to be tilted with respect to the CCD rows and columns, and even somewhat curved. However, these deviations can be compensated for during the calibration process.

5.3.7 Conversion to single-HDU FITS format

As explained above, one mosaiced RSS exposure is stored in a FITS file with 2 HDUs, a primary with a header and an extension with another header and the data

unit. This format is awkward to handle in further processing with IRAF, therefore the IRAF task **imcopy** was used to convert each exposure to a FITS file with only one HDU. The data unit of this HDU contains the data of the original extension, and its header was obtained by merging the headers of the original two HDUs.

5.3.8 Master flat creation

For those observations where flat frames were available with the same grating, grating angle and filter settings as the target data, the data were divided by a master flat (see §5.3.10). Just before the conversion to single-HDU format, the master flat was obtained in one of the following two ways: If raw flat frames were provided, they have already passed through the steps described in §5.3.2 to §5.3.6, and they were then combined with the **saltred** task **saltcombine** to yield a single master flat. Or, if only a master flat in the product folder was available, the master bias (if available) was subtracted from it just like from the raw frames, and **saltmosaic** was applied to it to obtain a single flat image. Such a master flat in the product directory had already undergone pre-processing, gain, cross-talk and overscan correction as part of the pipeline processing.

As mentioned in §5.3.3, the gain factors for the different amplifiers are often variable, and even after passing through **saltgain**, there may be slight gain differences between the amplifiers. These can be seen as “steps” of several counts between adjacent amplifiers when scrutinizing a flat field or the trace of a bright target. This was corrected with a PYTHON script¹⁰.

Consider one of the three CCD chips, consisting of two amplifier images. For each amplifier image, the median count value in an 11×5 region around the central row was determined. For the amplifier image on the left hand side, the region was chosen just to the left of the amplifier boundary, and for the one on the right, just to the right side of the boundary. The difference between the two median values then represented the “step” size between the two amplifiers caused by gain differences. The ratio between the medians was then used to obtain a multiplicative factor to correct for the residual gain difference between the two amplifiers.

This process was repeated for all three the CCD chips, and the derived multiplicative factors were applied to the master flat and the target and arc frames to correct the residual gain differences between the amplifiers. It should be noted that no correction was applied across the gaps *between* the CCDs, because in cases where there is a large gradient in the flat field, there *should* be a difference in counts to the left and right of a chip gap, regardless of gain errors.

¹⁰Written by the author, based on an IDL script of P. Vaisanen.

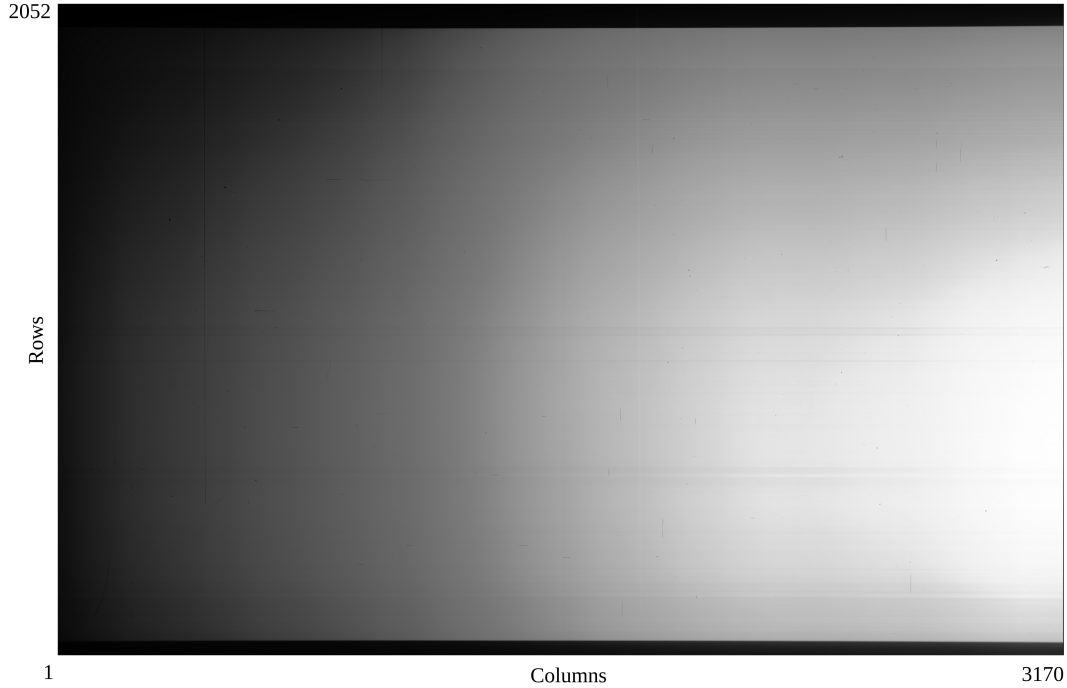


Fig. 5.4: RSS master flat obtained on 5 January 2013. The final gain correction have been performed and the gaps between the chips filled by interpolation, but no illumination correction has been performed yet.

The next step was to fill the chip gaps in the flat field. A PYTHON script was utilized to achieve this for each row independently, by fitting a straight line across the gap, using 9 pixels to the left and 9 pixels to the right of the gap (ignoring the column immediately adjacent to the gap on each side). The zero-count pixels in the chip gaps were then replaced by the corresponding fit value, yielding a suitable interpolation across the gaps. The master flat used for the CAL 83 observation on 5 January 2013 is shown in Fig. 5.4, after the final gain correction and “gap-filling”.

Unfortunately, the master flat is not ready to use yet. The difficulties involved in obtaining a flat frame that is really spatially and spectrally flat are discussed in §D.7. Indeed, the dramatic large-scale structure in the flat in Fig. 5.4 is mostly due to non-uniform field illumination by the flat lamp, as well as the spectral energy distribution of the flat lamp itself, yielding the large gradient in the dispersion direction (the lamp is brighter in the red than in the blue). It is true that any large-scale variations in the CCD sensitivity will also be present in the flat field, but without any additional information (e.g. blank sky flats) it is impossible to separate these from the lamp illumination signatures. Dividing the science frames by the flat field in this form will introduce artificial large-scale gradients into the data, which is highly undesirable.

This is avoided by first applying an illumination correction to the master flat, using the **mkillumflat** task within the IRAF package **imred.ccdred**. The first step performed by this task is to determine the illumination pattern by smoothing the flat field, using a moving two-dimensional “boxcar” average. The box size is not fixed, but rather increases from a specified minimum at the edges to a specified maximum at the centre of the image. In this case, the box size was set to vary from 3 to 150 pixels ($xboxmin = yboxmin = 3$, $xboxmax = yboxmax = 150$).

The illumination pattern therefore represents a heavily smoothed version of the original flat field. Consequently, **mkillumflat** creates an illumination corrected flat by dividing the original flat by the illumination pattern, yielding a master flat with the large-scale structure removed, but retaining the smaller scale and pixel-to-pixel variations. In this way, a compromise can be reached: applying this master flat will not remove large-scale structure in the CCD sensitivity, but will correct small-scale sensitivity variations in the CCD, as well as other small-scale structures, e.g. due to dust motes.

The illumination corrected flat was normalized by dividing all the pixels by the median pixel value, yielding the final master flat that will be used for processing the target and arc frames (§5.3.10).

5.3.9 Bad pixel map creation

The normalized master flat was used to create a bad pixel map to apply to the target and arc frames. A simple PYTHON routine was applied to flag all the pixels with a normalized value below 0.75 or above 1.25. The bad pixel map was stored as a regular FITS image, with pixel values of 0 indicating good pixels, and pixel values of 4 indicating bad pixels. A value of 4 was used because setting a value larger than 3 will yield interpolation across a bad pixel region along its narrowest dimension during processing with **ccdproc** (§5.3.10).

In some cases, flat frames with the same instrumental configuration as the target frames were not available, and no master flat division was applied to the science frames. However, in these cases a flat frame with the same binning as the target frames and taken close to the target observation in time, was still used to create a normalized flat (as described in §5.3.8), solely with the purpose of creating a bad pixel map. This is acceptable because the bad pixels are present in any flat at the same detector positions. Therefore a bad pixel map was applied to all observations, regardless of whether flat field division was performed.

5.3.10 Processing with `ccdproc`

The task `ccdproc` in the IRAF package `imred.ccdred` was used to perform the final reductions on the target and arc images. If applicable, these images were divided by the normalized master flat field, and the regions flagged in the bad pixel map were fixed by interpolating across their narrowest dimension.

Even though the overscan regions have been trimmed off in Fig. 5.3 and Fig. 5.4, one can see that there are still rows at the top and bottom without useful data. These can also be trimmed off. In fact, when one is only interested in the spectrum of a single target for each observation, only the rows containing this target’s trace, as well as additional rows for reasonable background determination need to be conserved. For every observation, only 50 rows below and 50 rows above the centre of the target trace were retained, with the rest of the rows trimmed off during `ccdproc` processing. This yielded target and arc images with dimensions 3170×101 .

5.3.11 Potential fringing corrections

The effect of fringing on our observations was not too severe, although it would have degraded the spectral quality on the far red side to some degree. Apart from the flat correction, no additional measures were taken to eliminate fringing. A very effective way of correcting the fringing effect would be to dither the target along the slit during the observations (e.g. SALT Ast Ops, 2014, p. 43). This simply means obtaining two or three target exposures, changing the target position on the slit in the spatial direction between each exposure. These target frames are then subtracted from each other in a pairwise fashion, and the final extracted spectra are then added together.

It is obvious that this subtraction method requires the target to be displaced along the slit, otherwise the subtraction would get rid of the target signal as well. With the target dithered along the slit, the subtraction eliminates the fringing pattern (which should be constant over the CCD image on short time-scales), and will also get rid of most of the background. For future SALT observations, we will make use of this strategy.

5.3.12 Stacking of spectral CCD images

During each observation, two or more consecutive exposures of the target were obtained with exactly the same instrumental configuration. After processing with `ccdproc`, these spectral images were stacked, i.e. combined to improve the signal-to-noise. This was accomplished with the IRAF task `imcombine`, which simply calculated the mean of the images. In cases where the exposures did not have identical exposure times, their contributions to the mean image were weighted by

exposure time, and scaled according to the inverse of the exposure time.

Some of the IRAF tasks that are used during the rest of the analysis require access to header keywords that indicate the Universal Time start date and time of the exposure, the local sidereal time at the start of the exposure, and also the exposure time. In the RSS files, these are stored in the header keywords DATE-OBS, UTC-OBS, LST-OBS and EXPTIME. However, when combining frames with **imcombine**, the keywords of the first frame are simply copied to the header of the combined frame, and need to be updated manually to realistic values for the combined exposure.

This update was performed with a PYTHON script utilizing several time and coordinate functions from the **astropy** module. First, the effective exposure time of the combined image was obtained by calculating the weighted mean of the exposure times of the individual frames. The centre of the whole observation was taken to be exactly between the start of the first exposure and the end of the last exposure, and the effective start time as the central time minus half of the effective exposure time. The header keywords DATE-OBS, UTC-OBS, LST-OBS and EXPTIME of the combined frame were then updated to reflect these “effective” values.

During the execution of **imcombine**, a cosmic ray rejection routine called “crreject” was enabled, which disposed of spurious pixels with very high counts relative to the surrounding pixels, thus yielding a combined frame with most of the cosmic rays removed.

5.4 Spectroscopic calibrations

After completion of the data reductions, the instrumental signature has been removed, and for each individual and combined exposure, one has a reduced two-dimensional image with pixels in $[x,y]$ coordinates, with the target trace extending over the image in the dispersion direction. The next task is now to transform these images to one-dimensional spectra, in the form of graphs of counts versus wavelength. This process involves several steps that are performed with tasks that can be found in the IRAF package **twodspec**. A detailed treatment of spectroscopic calibrations with IRAF is presented by Massey, Valdes and Barnes (1992).

5.4.1 Two-dimensional wavelength calibration

Wavelength calibration refers to the process during which the pixel scale in the dispersion direction is translated to a wavelength scale, i.e. values in Å. This is achieved by taking comparison arcs during observations. A comparison arc is a (usually short) exposure of a discharge tube, and such an image is characterized by a number of



Fig. 5.5: Trimmed argon comparison arc obtained with the RSS on 5 January 2013. Note that the arc lamp emission lines extend over the whole image in the spatial direction.

sharp emission lines extending over the whole slit in the spatial direction. The laboratory wavelengths of the emission lines in these discharge tubes are very well known, and the line wavelength lists (“line atlases”) for all the available arc lamps used with the RSS, as well as plots of the arc spectra, can be found on the PySALT website¹¹.

The image of an argon arc lamp obtained for the wavelength calibration of the CAL 83 spectra of 5 Jan 2013 is shown in Fig. 5.5. Note that, at this stage of the calibration, the target and arc spectra have been trimmed to a size of 3170×101 , therefore the spatial extent is much smaller than in Fig. 5.3 and Fig. 5.4. The central cross-section of the Ar arc in Fig. 5.5 is shown in Fig. 5.6, consisting of the average of rows 50, 51 and 52.

The task **identify** in the IRAF package **twodspec.longslit** was used to assign wavelengths to the arc emission features in Fig. 5.6. This was accomplished by comparing the arc spectrum at hand with the known arc spectrum plot, and identifying four or five prominent emission features distributed throughout the spectrum. The wavelengths for these features were provided as input to **identify**, and the task fitted a sixth-order legendre polynomial function to provide a preliminary transformation function from x -pixel to wavelength. The task then transformed the arc spectrum to wavelength coordinates, and automatically identified the other emission features by looking for them at the expected feature positions given in the relevant line atlas (Ar in this case).

A final fit was then performed by **identify** using all the feature positions, and applied to the arc spectrum (as shown in Fig. 5.6). The task then stored a file containing the x -pixel positions and corresponding wavelengths of the features in Å, as well as the fit parameters providing the x -to- λ transformation. However, it is quite evident from Fig. 5.3 that the sky emission features are not completely vertical, but rather curve slightly over several columns due to distortion in the system. The same effect is obvious in the arc frames before trimming. After trimming (e.g. in Fig. 5.5) the extreme ends of the curves are trimmed off, but the effect is still present. The x -to- λ solution along the central row will therefore not be valid for every row. To obtain a two-dimensional wavelength solution, one needs to repeat the identification process

¹¹<http://pysalt.salt.ac.za/lineatlas/>

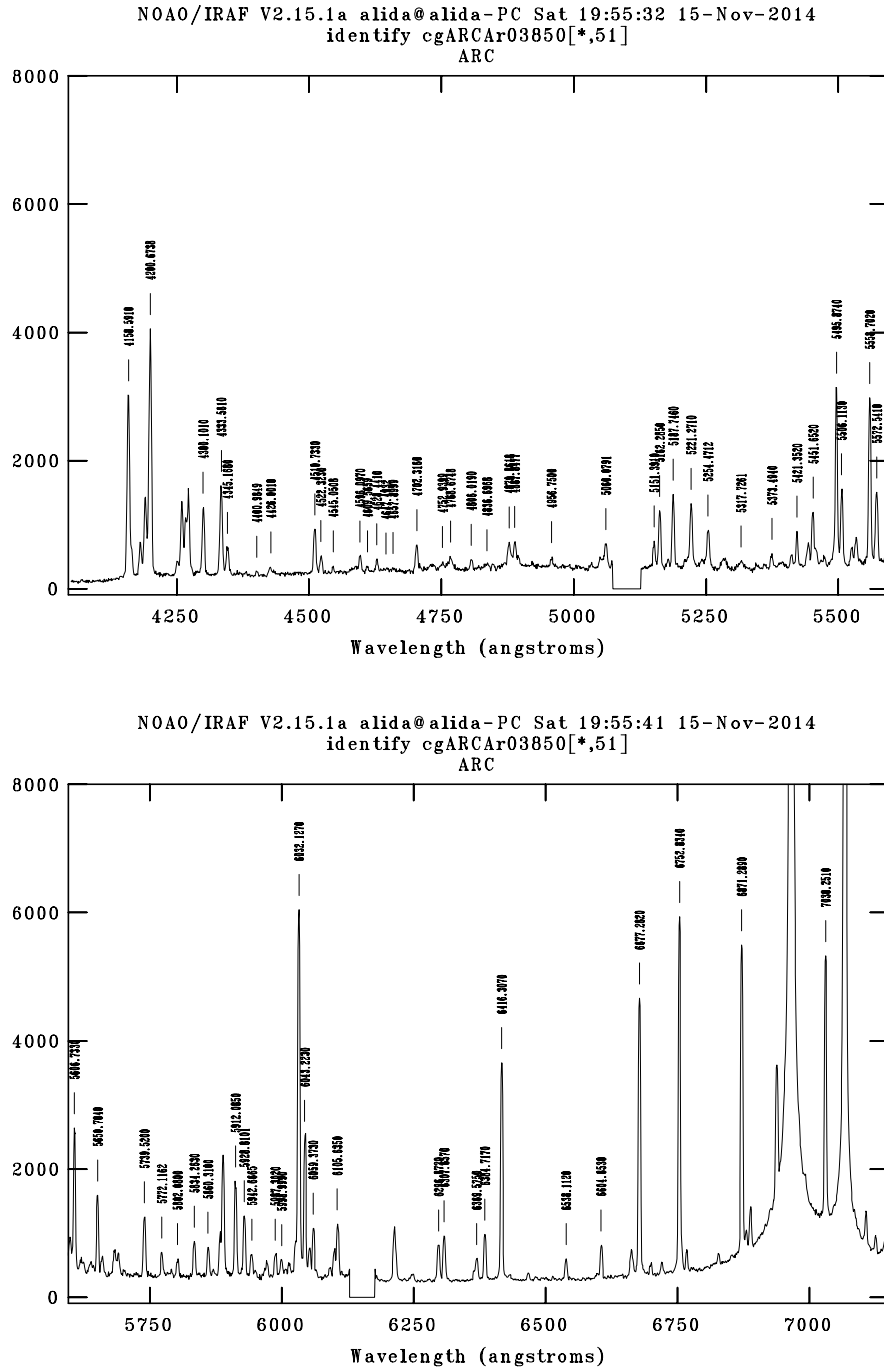


Fig. 5.6: The cross-section of the Ar arc frame in Fig. 5.5, extending from blue (top) to red (bottom), obtained by averaging 3 rows centred on row 51. The identified arc lines are indicated, and the horizontal axis has already been transformed wavelength coordinates.

across the whole two-dimensional arc image.

This was achieved with the task **reidentify**, which looked for the same comparison features found by **identify** in every second row of the CCD image. For each of these rows, the mean of 3 adjacent rows were used, thus providing a moving average of the feature positions across the image. The shape of the fit function determined by **identify** was not changed by **reidentify**, but for each row under consideration the latter only determined the shift of the solution in the dispersion direction relative to the solution of the central row. After this process, one has a file containing the $[x,y]$ positions of arc features over the whole CCD image, together with the wavelength associated with each point.

These data points were provided as input to the **fitcoords** task, which fitted a two-dimensional legendre polynomial function (with order 6 in the x -direction and order 4 in the y -direction) to the points. The result was a surface function enabling the transformation of an image with $[x,y]$ coordinates to an image with $[\lambda,y]$ coordinates. Finally, using the task **transform**, this function was applied to the individual target frames, the combined target frames, and also to the arc frames themselves to monitor the validity of the transformation. After the transformation, the arc lines were completely straight, and extended vertically over a “ λ -column”. The same applied to the sky lines in the transformed target exposures, enabling accurate background subtraction, as discussed in the next section.

It is very important that the wavelength solution be very accurate. In other words, the wavelength solution determined from the arc frame must be just as valid for the target frames it is applied to as for the arc frame itself. Therefore, the arc frame must be recorded with *exactly* the same instrumental configuration as the target frames, and *directly* before or after the target frames. For most of the RSS observations reported here, the arcs were taken just after the target frames, or in some cases just before.

Of course, the wavelength solution can not be completely perfect, and the typical rms error applicable to the two-dimensional function produced by **fitcoords** and applied by **transform** is given in Table 5.4 for each configuration, serving as an indication of the uncertainty associated with the wavelength calibration. These rms values range between 0.2 and 0.8 Å.

5.4.2 Background subtraction

Two-dimensional background subtraction was performed using the **background** task. This task fitted a cubic spline consisting of 3 pieces to the background in

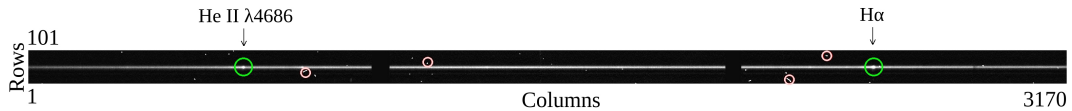


Fig. 5.7: Spectral CCD image of CAL 83 obtained with the RSS on 5 January 2013 (first exposure). This is the same frame that was shown in Fig. 5.3, but with all processing including background subtraction completed. Note the absence of background emission lines. The positions of the strongest emission lines in CAL 83 are indicated by the large green circles and the arrows. A few cosmic rays in the image have been marked with the smaller circles.

every λ -column, and subtracted this fit from the column, eliminating the sky and other extended background emission and leaving only the target trace. The target trace was ignored during the fit by utilizing a rejection algorithm to reject pixels with high counts, to avoid subtracting any emission from the target itself. This ignored cosmic rays as well, and a lower limit was also used to reject dead pixels. For the data at hand, it was found that the target trace was appropriately ignored while still fitting the background accurately by rejecting points above 1.5 times the residual sigma, and points below 4 times the residual sigma, using 5 to 7 rejection iterations.

The processed, trimmed, transformed and background-subtracted version of the CAL 83 frame in Fig. 5.3 is shown in Fig. 5.7. It is clear that the sky lines have been removed, but the cosmic rays are still present. Recall that in §5.3.12, only cosmic rays in the combined images were removed. For the individual exposures, the cosmic rays were removed during the spectral extraction described in the next section. The broader bright spots superimposed on the target trace represent emission lines in the target spectrum.

5.4.3 Extraction of the one-dimensional spectra

The task **apall** in the **twodspec.apextract** package was utilized to define an aperture containing the target trace on the two-dimensional target image. In general, the target aperture was defined to extend from 10 rows below the centre of its trace to 10 rows above the trace, i.e. a total height of 5.2 arcsec. Care must be taken not to include the trace of any other source along the slit in the aperture. In the example spectrum, another trace was present 20 rows below the CAL 83 trace, well outside the aperture.

Due to distortion, the target trace is usually not completely “horizontal” (aligned with the rows). Instead, the trace can be slightly tilted, cutting over several rows from left to right, and can even be somewhat curved. To account for this, the “tracing” capability of **apall** was activated for most of the spectra, which tracked the target trace across the image, and applied a similar distortion to the rectangular

aperture to make sure that the target trace centre is always in the centre of the aperture. However, if some parts of the trace are very faint, the process is bound not to be very successful, and in these cases a simple rectangular aperture that included the whole trace was used, with an increased height of 7.8 arcsec if necessary (extending from 15 rows below to 15 above the trace centre).

Finally, **apall** extracted the spectrum from the region in the aperture by summing the counts along each λ -column. Variance weighting was applied, i.e. larger weight assigned to the higher counts values closer to the central row. The cosmic ray cleaning algorithm in **apall** was also activated to eliminate the cosmic rays in the individual frames, and also any residual cosmic rays left in the combined frames. The result was a single count value associated with each wavelength value, i.e. a one-dimensional spectrum of counts versus wavelength.

5.4.4 Additional header updates

A few final updates to the headers of the individual and combined target spectra were performed to facilitate the ensuing analysis process. Firstly, the value of the OBSERVAT keyword was changed from “SALT” to “SAAO” to enable recognition by the IRAF tasks. Then, the existing AIRMASS keyword was renamed to OLD_AIRM, and an updated value calculated with the **onedspec.setairmass** task was stored under AIRMASS. This task also calculated the Universal Time at the middle of the exposure, and stored it under the UT keyword.

The **onedspec.setjd** task was utilized to calculate the Julian day (JD), heliocentric Julian day (HJD) and local Julian day number associated with the middle of the observation, and to write these values to the header of each frame. The Barycentric Julian Date (BJD) in the Barycentric Dynamical Time standard (TDB) at the middle of the observation was determined by using the ‘UTC2BJD’ Time Utility of the Ohio State University¹² (Eastman et al., 2010), and also stored in the header.

The task **bcvcorr** in the **rvsao** package was applied to calculate the barycentric velocity correction for each spectrum, i.e. the speed of the observatory towards the solar system barycentre (SSB) projected on the line of sight to the target at the time of observation. This value was stored in the BCV header keyword, and utilized in §5.5.5 to calculate the radial velocity of the target relative to the SSB.

5.4.5 Finalization of spectra

The task **spectool** forms part of the **guiapps.spt** package, and can be used for spectral viewing and analysis. As a final check, each spectrum was loaded into

¹²<http://astrutils.astronomy.ohio-state.edu/time>

spectool and compared with the other spectra for the same target and day, as well as with its corresponding spectral CCD image, to confirm that all cosmic rays and other unwanted features have been removed during the calibration process.

Overplotting of all the spectra for a given target and date also showed that there were no apparent changes in any of the spectra from one exposure to the next, therefore the subsequent analysis was carried out only for the combined spectrum for each day. The exception is the two CAL 83 datasets of 21 March 2013, which were taken at slightly different grating angles and therefore not combined. Finally, each FITS file containing a target spectrum was converted to a text file containing only the wavelength and count columns with the task **onedspec.wspectext**. This was necessary to input the data to the Fityk software.

5.5 Spectral analysis

After completing all the calibrations, one comes to the most important part of the process: extracting astrophysically useful information from the spectra. Modelling of the continuum shape did not form part of the current analysis, and no absolute flux calibration or continuum shape calibration with standard stars was performed. The focus of the analysis reported here was on the emission lines, specifically determining the following for each line: line identity, full width at half maximum, full width at zero intensity, equivalent width, measured central wavelength and the associated barycentrically corrected velocity shift from the laboratory wavelength.

5.5.1 Identification of spectral lines

First of all, all the spectral lines in the spectra were identified. This was achieved by comparing the new spectra to previously published spectra of the respective sources (see Chapter 3), as well as similar sources, and also by consulting the NIST Atomic Spectra Database¹³ (Kramida et al., 2014). For the wavelengths of the ionized helium lines, see the summary of Varshni and Talbot (2006) and references therein.

5.5.2 Defining the continuum

After loading the spectrum into the Fityk software package, the first step was to define the continuum. This was performed with great care for each spectrum, as an accurate continuum level is very important for measuring the correct parameters for the spectral lines. The continuum level was estimated by eye in Fityk by defining a point approximately every 100 Å, and using a cubic spline interpolation function between the points. The CAL 83 spectrum of 5 January 2013 with its defined continuum (known as the “baseline” in Fityk) is shown in Fig. 5.8. The continuum

¹³http://physics.nist.gov/PhysRefData/ASD/lines_form.html

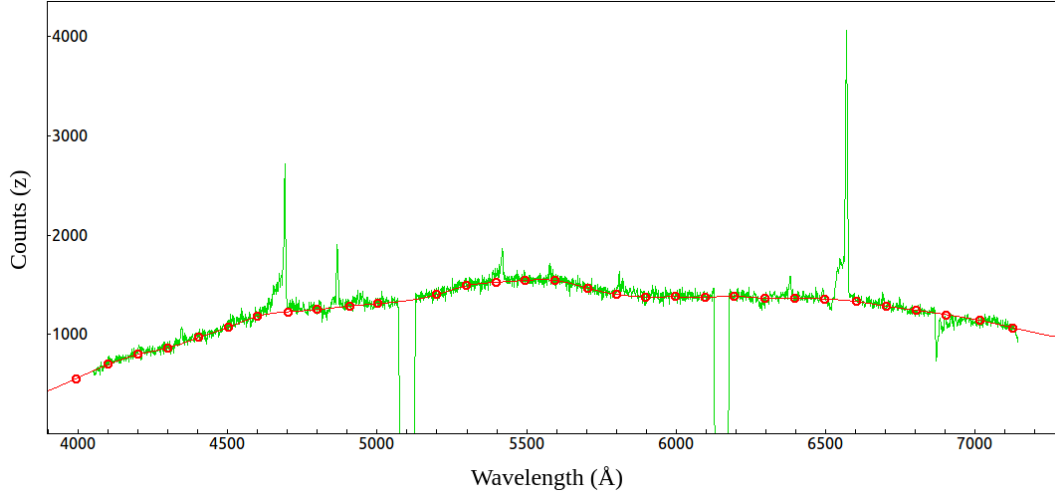


Fig. 5.8: The combined RSS spectrum of CAL 83 obtained on 5 January 2013 (green), together with the continuum defined in Fityk (red). The open red circles represent the points defined by the author, and the red lines between the cubic spline interpolation.

function was subsequently subtracted from the spectrum.

In Fityk, the counts values of the data points before and after continuum subtraction are designated by the parameters y and Y respectively. However, to avoid confusion with the y -dimension of the CCD images described before, the notation z and Z will be used to indicate counts before and after continuum subtraction.

5.5.3 The mean error and signal-to-noise

For reasonable error estimations, it is important to have knowledge about the typical fluctuations in the spectrum due to noise, as well as the signal-to-noise ratio (S/N) of the spectrum. To this end, three or four regions that did not contain any apparent emission or absorption lines were identified throughout the spectrum. The standard deviation of the continuum-subtracted data in each of these regions were determined with the `stddev` command in Fityk. The mean of these standard deviations was determined, and this was considered as the typical counts error in the spectrum.

Then, the mean value of the continuum function in the same wavelength ranges were determined with the `avg` command. Dividing this value by the standard deviation determined above provided a robust estimate of the S/N of the spectrum.

5.5.4 Fitting spectral lines with Fityk

The line profile fitting was also performed with the Fityk software package, utilizing the Levenberg-Marquardt fitting method. A simple Gaussian profile was fitted to emission lines with a simple single-peaked profile. In cases where the spectral line consisted of multiple components, multiple Gaussian profiles were fitted simultane-

ously. The process described in this section will be with reference to a single fitted Gaussian profile, but note that in cases where more than one profile were fitted simultaneously, the process was applied to each component.

In Fityk, a Gaussian profile is characterized by three parameters. The first is the “height” of the line, i.e. the number of counts that the line peak is above the continuum level, which will be referred to here as Z_{max} . Secondly there is the line “center”, i.e. the wavelength position of the peak, which will be called λ_{obs} . The last parameter is the “hwhm” (half width at half maximum) of the line in Å, which will be referred to as HWHM_{obs} . After obtaining the best fit parameters, the observed HWHM was corrected for instrumental broadening by using Eq. (C.9). The “hwhm” parameter of the fit was then set to the corrected value $\text{HWHM}_{\text{corr}}$.

The confidence intervals for the fitted parameters can be derived in Fityk for any desired confidence level. As the measured parameters depend on the accurate continuum definition, which will include some amount of uncertainty, it was decided to record 3σ (99.73%) errors for Z_{max} , λ_{obs} and $\text{HWHM}_{\text{corr}}$, rather than 1σ (68.27%) errors, to be on the conservative side.

The next step was to determine the wing limits of the line profile, i.e. the wavelength positions where the wings blend into the continuum. As the fitted profile will approach the defined continuum asymptotically on each side, it was decided to choose the wing limits at the positions λ_b on the left and λ_r on the right where the profile intensity has decreased to 0.1% of the peak height Z_{max} . For a Gaussian profile, 99.9% of the area under the profile is contained within these limits, enabling an accurate equivalent width calculation.

Subsequently, the following parameters were determined between the limits λ_b and λ_r : the mean counts in the line, $\bar{Z}_{\text{line}}$, the mean continuum level $\bar{z}_c$, and the area between the line and the continuum, A_{line} (using the **darea** command for the latter).

5.5.5 Final spectral line calculations

The corrected full width at half maximum $\text{FWHM}_{\text{corr}}$ was calculated from $\text{HWHM}_{\text{corr}}$ with Eq. (C.6). The observed velocity of the line (v_{obs}) was determined with Eq. (C.1), and the velocity referred to the SSB (v_{SSB}) with Eq. (C.2), after reading the appropriate value for BCV from the spectrum header. The equivalent width was obtained with

$$W_\lambda = \frac{-A_{\text{line}}}{\bar{z}_c} , \quad (5.2)$$

which is just another form of Eq. (C.7), assuming that the continuum level between λ_b and λ_r has a constant value of $\bar{z}_c$. Eq. (C.8) was utilized to determine the error in

the equivalent width, with $\overline{F}_c = \overline{z}_c$ the mean continuum level and $\overline{F} = \overline{z}_c + \overline{z}_{\text{line}}$ the mean counts in the line before continuum subtraction. Upon obtaining the equivalent widths of all the emission lines, a selection of relevant line ratios were also calculated for each source.

Now that the reduction and analysis techniques have been reviewed, the spectra of each of the sources observed with SALT (see §5.2) will be presented in the following sections.

5.6 CAL 83

5.6.1 Results

The calibrated, continuum-subtracted spectra of CAL 83 obtained during the observations listed in Table 5.3 are shown in Fig. 5.9, with the detected emission lines marked on each spectrum. The properties of all the emission lines that have been detected previously in the central source in CAL 83 are listed in Table 5.5 (see §3.1.3). For reference, their positions are also indicated in the spectra in Fig. 5.9, regardless of whether or not they were observed in the respective spectra. The line labels in a particular spectrum were offset by a fixed velocity equal to the mean velocity of the fitted lines in that particular spectrum. It was also noted that the two-dimensional spectrum before background subtraction seemed to exhibit the [O III] nebular emission lines at 4959, 5007 and 6302 Å, which warrants further investigation as part of follow-up work.

The most prominent features in the spectra are the H α and He II λ 4686 emission lines. Emission lines of H β , H γ and H δ are also visible in most spectra (the latter only very faintly), although they were not detected in the low S/N spectra on 11 Sep 2012, 29 Oct 2012 and 21 Mar 2013. Apart from He II λ 4686, several other lines from the He II Pickering series are also evident: He II λ 4200, He II λ 4541 and He II λ 5411. Although previous publications have classified the feature around $\sim$ 6680 Å as He I λ 6678, it may well be He II λ 6683, or possibly a blend between the two, as other He I lines, e.g. He I λ 5876, are not detected in the spectra. The H α and H β lines are expected to be blended with He II λ 6560 and He II λ 4859 respectively.

The Balmer and He II lines exhibit broad blue wing structures extending to $\sim$ 2000 km s $^{-1}$. To better illustrate these broad structures, the H β , H α and He II λ 4686,5411 lines observed on 5 Jan 2013 are plotted in terms of velocity relative to the central line wavelength in Fig. 5.10. This figure also illustrates the presence of P Cyg blueshifted absorption structures at $\gtrsim$ 2000 km s $^{-1}$ in the Balmer lines.

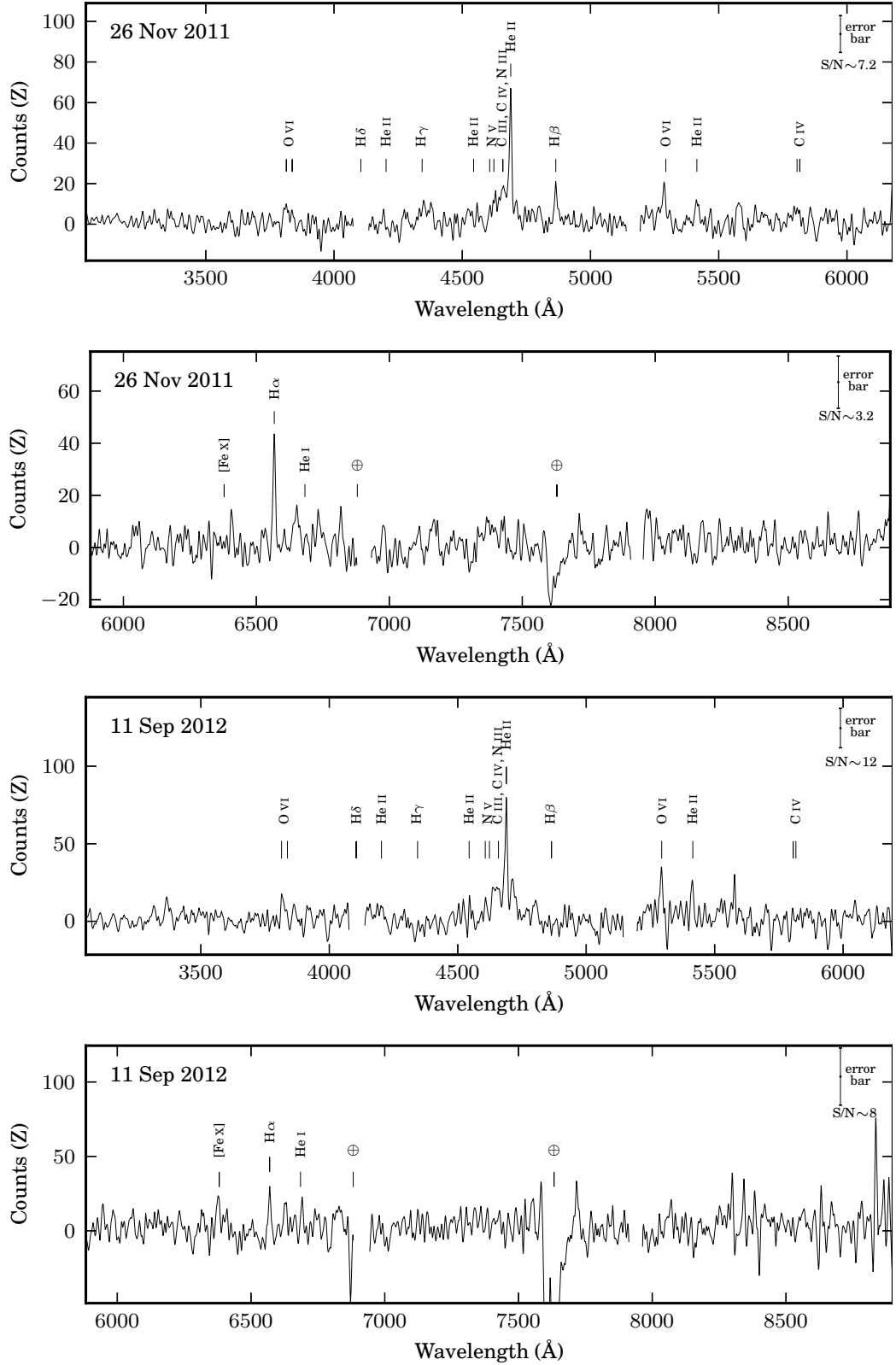


Fig. 5.9: SALT RSS spectra of CAL 83. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

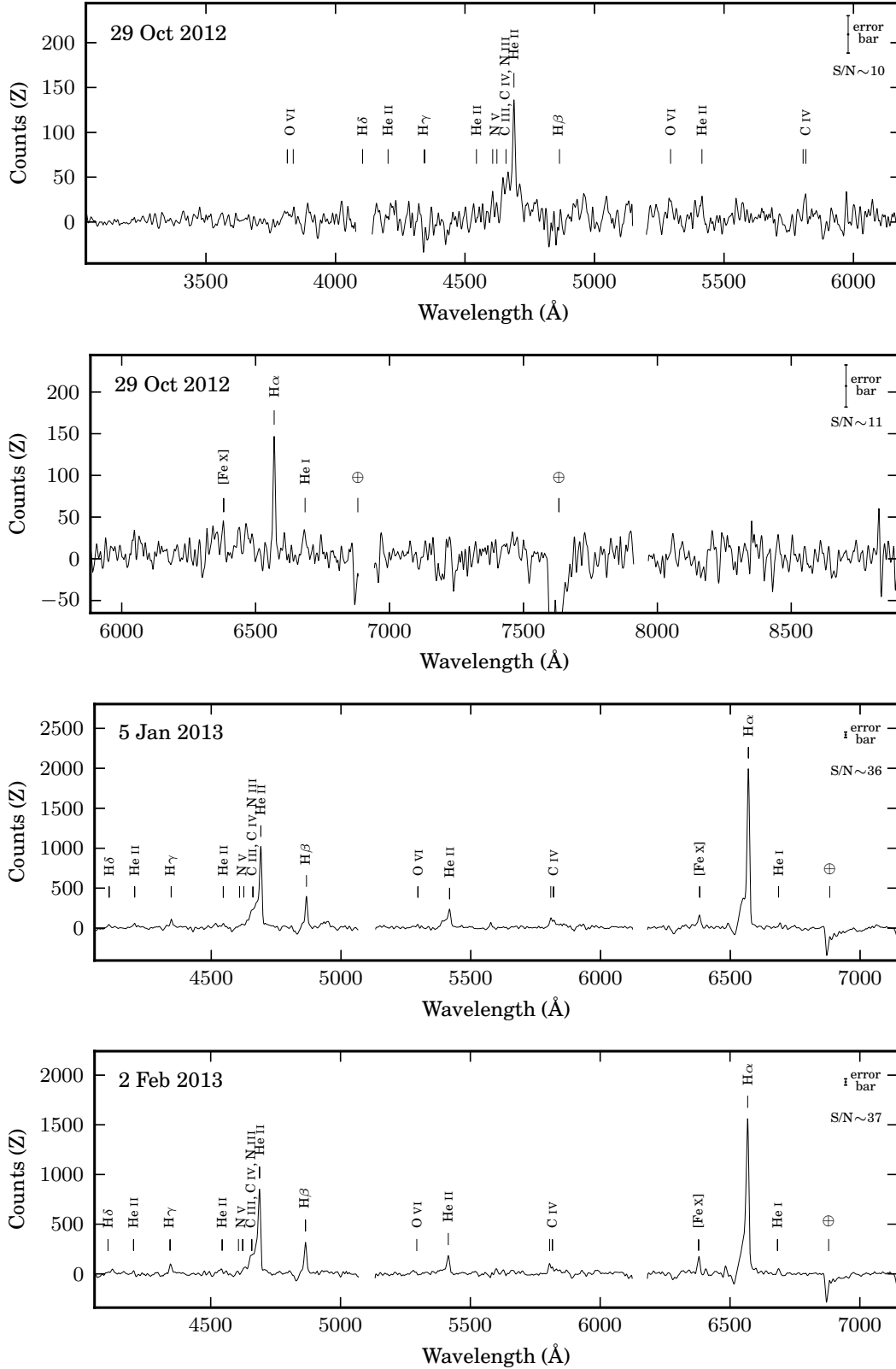


Fig. 5.9: (continued) SALT RSS spectra of CAL 83. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

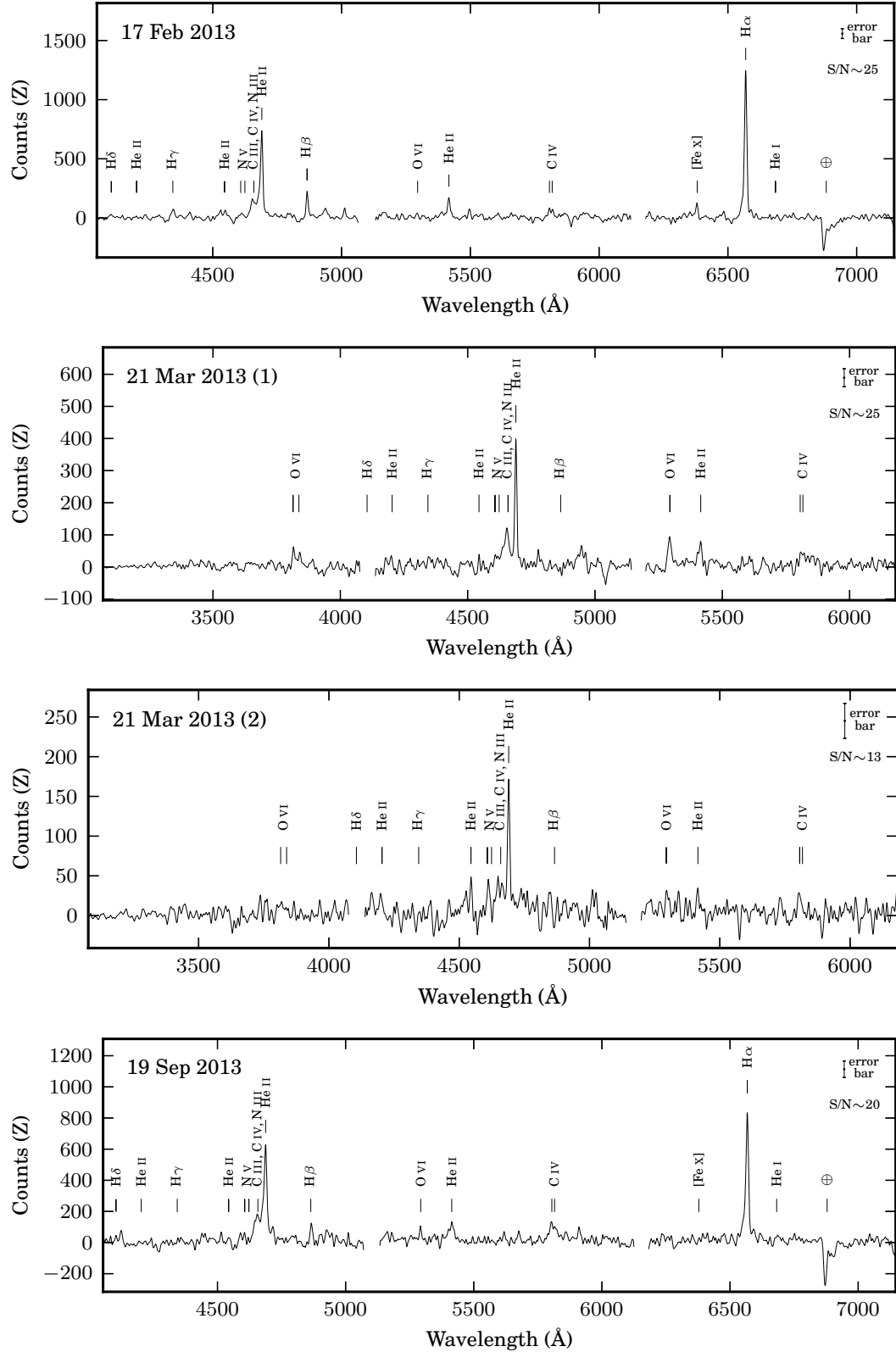


Fig. 5.9: (continued) SALT RSS spectra of CAL 83. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

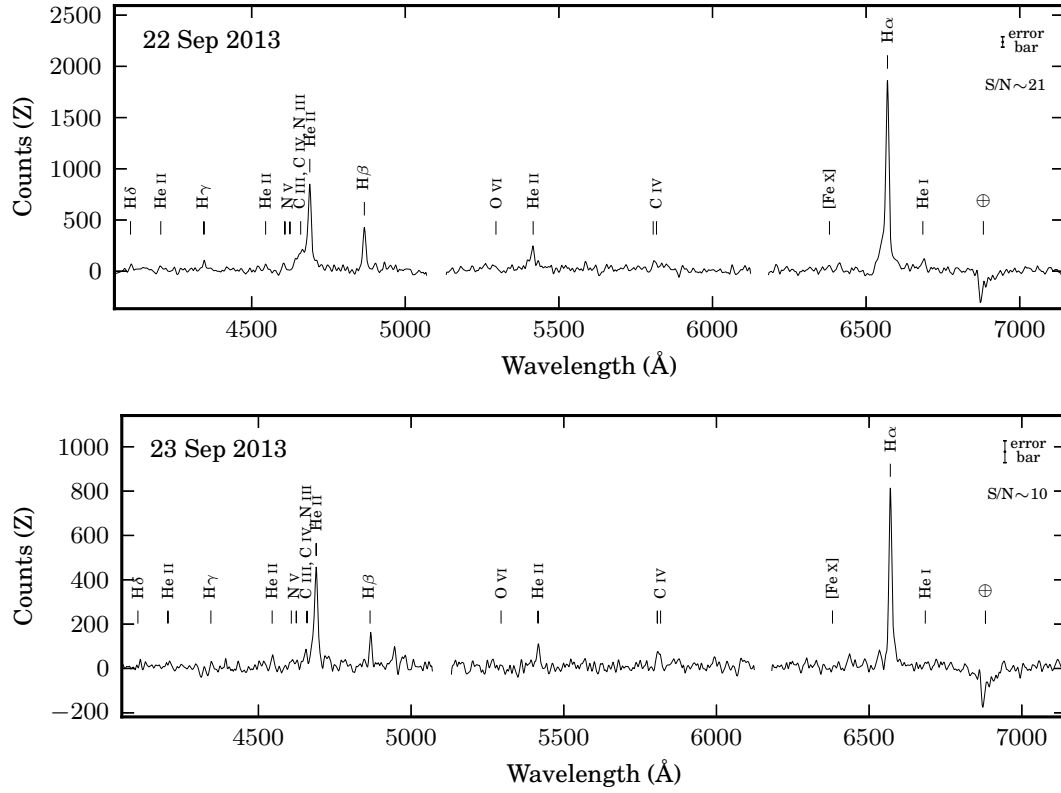


Fig. 5.9: (continued) SALT RSS spectra of CAL 83. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

Table 5.5: Standard properties of the emission lines observed in CAL 83. The third column indicates the temperature range within which the ion can exist in the relevant ionization state without being excited to an even higher state, according to the ionization potentials and assuming thermal ionization. The last column represents the thermal broadening FWHM according to Eq. (C.11) for the single lines, corresponding to the highest temperature in the third column. The rest wavelengths and ionization potentials were obtained from the NIST Atomic Spectra Database.

Element and ionization state	Rest wavelength λ_0 (Å)	Temperature range (K)	Thermal broadening FWHM (Å)
H I (H α) (Blended with He II λ 6560)	6562.79	<157700	1.9
H I (H β) (Blended with He II λ 4859)	4861.35	<157700	1.4
H I (H γ)	4340.472	<157700	1.2
H I (H δ)	4101.734	<157700	1.2
He I	6678.151	<285200	1.3
He II	4200	285200-631200	1.2
He II	4541	285200-631200	1.3
He II	4686	285200-631200	1.3
He II	5411	285200-631200	1.5
Blend of C III C IV, and N III emission	~4630-4680	282900-555500 555500-748100 343400-550400	
C IV	5801.33	555500-748100	1.0
C IV	5811.98	555500-748100	1.0
N V (?)	4603.73	898700-1136000	0.9
N V (?)	4619.98	898700-1136000	0.9
O VI	3811.35	1321000-1602000	0.9
O VI	3834.24	1321000-1602000	0.9
O VI	5290	1321000-1602000	1.2
[Fe X]	6374.51	2710000-3040000	1.1

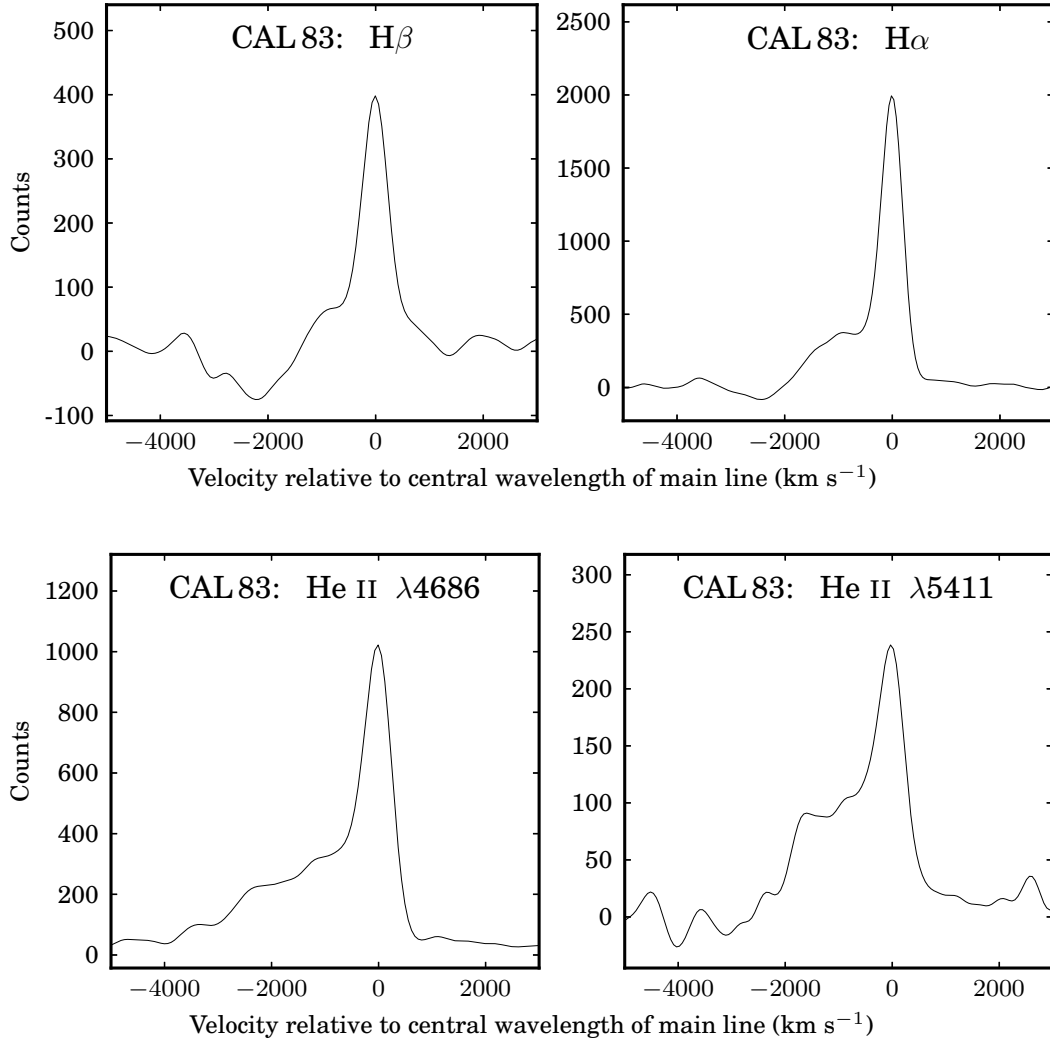


Fig. 5.10: The spectral regions around the H β , H α and He II $\lambda\lambda$ 4686,5411 emission lines in the CAL 83 spectrum on 5 Jan 2013, plotted in terms of velocity relative to the central line position. Note the broad, blueshifted bases of the lines, as well as the blueshifted absorption features in the Balmer lines. C III, C IV and N III emission is blended with the He II λ 4686 wing.

Emission of highly ionized species are also present in the spectra, i.e. O VI $\lambda\lambda 3811, 3834$, faint traces of N V $\lambda\lambda 4602, 4620$, the combination of C III, C IV and N III (blended with the blue wing of He II $\lambda 4686$), O VI $\lambda 5290$, C IV $\lambda\lambda 5801, 5912$, as well as [Fe X] $\lambda 6375$.

Detailed fitting as described in §5.5.4 and §5.5.5 was only performed for the strongest Balmer and He II lines in each spectrum, as well as for the O VI $\lambda 5290$ and [Fe X] $\lambda 6375$ lines where the signal was high enough for a reasonable fit to be obtained. The results are presented in Tables 5.6, 5.7 and 5.8.

The H α and H β lines on 5 Jan 2013 and 2 Feb 2013 were fitted with a combination of three gaussian profiles: one for the main emission component, one for the broad wing structure, and one for the blueshifted absorption component (e.g. Fig. 5.11). In the spectra obtained on 17 Feb 2013 and during Sep 2013, the blueshifted H α absorption was not significant enough to perform a reasonable fit, and only two Gaussian components were used: one for the main emission and one for the broad wing. For these spectra, only the main H β emission could be fitted, as the wing structure was not discernible above the noise level. For the 2011 and 2012 spectra, only the main component of H α was fitted (with a single Gaussian profile), and no fits were performed to the H β emission due to the low S/N. The strength of H γ emission was only adequate for a reasonable fit in the Jan and Feb 2013 spectra.

The main He II $\lambda 4686$ emission line was fitted with a Gaussian profile, with an additional component for the wing structure (e.g. Fig. 5.12). When the intensity of the blue wing is considerable, it blends with the C III, C IV, N III complex at ~ 4650 Å, and they can not be distinguished. Therefore, the Gaussian wing component should be considered as a combination of the wing structure and the blend of ionized carbon and nitrogen emission. The exception is the spectrum of 17 February 2013, where the wing structure was weaker, and a combination of three Gaussian profiles were fit to the data: one for the main emission component, one for the wing structure and one for the Bowen blend (see Fig. 5.13).

The He II $\lambda 5411$ main emission line and its broad wing component were fitted for the Jan and Feb 2013 observations, as well as the one on 22 Sep 2013. The [Fe X] $\lambda 6375$ emission line was fitted for the Jan and Feb 2013 spectra, while the O VI $\lambda 5290$ emission line was fitted for the spectra obtained on 26 Nov 2011, 11 Sep 2012 and 21 Mar 2013.

By utilizing the measured equivalent widths (W_λ), a selection of line ratios were calculated for each spectrum. No variability in the value of any of the line ratios was detected outside the 99.73% error margins. Consequently, the mean of each line

Table 5.6: CAL 83 Balmer emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing and absorption components it is with respect to the central wavelength of the relevant main component.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
Hα main component					
26 Nov 2011	5891.58284	202 ± 38	6.7 ± 2.0	22.9	−12 ± 14
11 Sep 2012	6181.62023	309 ± 86	5.1 ± 4.5	17.1	−1.5 ± 3.1
29 Oct 2012	6229.57263	289 ± 29	5.8 ± 1.5	20.0	−5.0 ± 3.1
5 Jan 2013	6298.49765	311.9 ± 4.2	6.19 ± 0.27	20.5	−12.1 ± 1.1
2 Feb 2013	6326.30355	242 ± 11	7.03 ± 0.76	22.5	−10.2 ± 1.1
17 Feb 2013	6341.38141	288.6 ± 6.7	5.73 ± 0.44	19.5	−8.5 ± 1.4
19 Sep 2013	6554.58238	204 ± 15	6.01 ± 0.91	20.5	−5.5 ± 1.8
22 Sep 2013	6557.63127	307.4 ± 8.0	6.37 ± 0.49	21.5	−11.9 ± 2.0
23 Sep 2013	6558.60809	300 ± 13	6.22 ± 0.78	20.5	−12.4 ± 4.2
Hα wing					
5 Jan 2013	6298.49765	−785 ± 91	34.5 ± 4.9	110.4	−10.7 ± 4.6
2 Feb 2013	6326.30355	−520 ± 140	28.6 ± 4.8	91.8	−10.6 ± 3.8
17 Feb 2013	6341.38141	−210 ± 130	26.8 ± 6.4	86.0	−4.7 ± 5.0
19 Sep 2013	6554.58238	−230 ± 200	39 ± 11	124.9	−6.8 ± 9.4
22 Sep 2013	6557.63127	−260 ± 100	40.3 ± 5.6	128.8	−14.5 ± 9.4
23 Sep 2013	6558.60809	120 ± 430	58 ± 27	183.5	−10 ± 27
Hα blueshifted absorption					
5 Jan 2013	6298.49765	−2360 ± 170	16.7 ± 8.5	53.7	1.5 ± 2.1
2 Feb 2013	6326.30355	−2430 ± 150	9.2 ± 7.9	29.3	1.0 ± 1.1
Hβ main component					
5 Jan 2013	6298.49765	339 ± 20	3.82 ± 0.88	13.7	−1.76 ± 0.59
2 Feb 2013	6326.30355	238 ± 29	4.8 ± 1.6	15.6	−1.61 ± 0.64
17 Feb 2013	6341.38141	289 ± 30	3.0 ± 1.1	10.8	−1.2 ± 0.65
22 Sep 2013	6557.63127	336 ± 41	7.3 ± 1.6	24.4	−4.4 ± 1.9
23 Sep 2013	6558.60809	365 ± 88	7.1 ± 3.4	23.4	−3.7 ± 3.8
Hβ wing					
5 Jan 2013	6298.49765	−1500 ± 1200	50 ± 23	159.2	−5.1 ± 6.4
2 Feb 2013	6326.30355	−120 ± 270	22 ± 13	69.3	−1.7 ± 2.7
Hβ blueshifted absorption					
5 Jan 2013	6298.49765	−2120 ± 330	28 ± 19	87.9	4.2 ± 3.3
2 Feb 2013	6326.30355	−2160 ± 180	12.9 ± 7.1	42.0	1.0 ± 1.6
Hγ main component					
5 Jan 2013	6298.49765	358 ± 76	5.4 ± 2.6	18.6	−1.05 ± 0.76
2 Feb 2013	6326.30355	330 ± 130	9.0 ± 4.3	29.3	−1.4 ± 1.2
17 Feb 2013	6341.38141	370 ± 200	12.0 ± 6.7	39.1	−1.4 ± 2.2

Table 5.7: CAL 83 He II emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing components it is with respect to the central wavelength of the relevant main component. In cases where the He II $\lambda 4686$ wing and C III, C IV, N III blend were fitted separately, two measurements are given. The first is applicable to the wing, and the second to the carbon-nitrogen blend.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
He II $\lambda 4686$ main component					
26 Nov 2011	5891.56587	207 ± 39	4.5 ± 1.5	14.9	−5.0 ± 3.7
11 Sep 2012	6181.60840	217 ± 30	2.8 ± 1.1	9.9	−2.0 ± 1.3
29 Oct 2012	6229.52166	209 ± 67	5.3 ± 2.6	17.9	−3.0 ± 2.9
5 Jan 2013	6298.49765	344 ± 12	4.15 ± 0.5	13.7	−4.65 ± 0.67
2 Feb 2013	6326.30355	162 ± 17	5.99 ± 0.75	19.5	−5.55 ± 0.90
17 Feb 2013	6341.38141	304 ± 16	3.56 ± 0.7	12.7	−3.39 ± 0.85
21 Mar 2013 (1)	6373.28286	238 ± 19	4.26 ± 0.75	14.7	−2.86 ± 0.94
21 Mar 2013 (2)	6373.30167	272 ± 42	3.5 ± 1.6	12.7	−2.5 ± 1.6
19 Sep 2013	6554.58238	137 ± 38	6.3 ± 1.5	20.5	−4.2 ± 1.7
22 Sep 2013	6557.63127	227 ± 22	5.6 ± 0.88	18.5	−5.7 ± 1.5
23 Sep 2013	6558.60809	220 ± 38	6.5 ± 1.5	21.5	−7.8 ± 3.9
He II $\lambda 4686$ wing + C III, C IV, N III blend					
26 Nov 2011	5891.56587	−1580 ± 930	84 ± 32	265.1	−17 ± 55
11 Sep 2012	6181.60840	−780 ± 960	102 ± 38	324.1	−12 ± 39
29 Oct 2012	6229.52166	−800 ± 1000	99 ± 39	311.9	−16 ± 46
5 Jan 2013	6298.49765	−1190 ± 170	46.6 ± 5.5	148.5	−12.9 ± 6.2
2 Feb 2013	6326.30355	−1120 ± 200	41.2 ± 6.1	130.8	−10.5 ± 5.3
17 Feb 2013	6341.38141	−430 ± 210	24.0 ± 6.6	76.2	−6.1 ± 4.5
	+	−2310 ± 210	17.6 ± 7.5	55.7	−3.0 ± 3.2
21 Mar 2013 (1)	6373.28286	−2420 ± 420	49 ± 16	155.2	−4.7 ± 8.9
21 Mar 2013 (2)	6373.30167	−700 ± 1900	133 ± 82	420.3	−12 ± 48
19 Sep 2013	6554.58238	−920 ± 420	63 ± 14	200.1	−13 ± 15
22 Sep 2013	6557.63127	−800 ± 300	63 ± 10	201.0	−16 ± 14
23 Sep 2013	6558.60809	−500 ± 790	79 ± 31	249.9	−15 ± 38
He II $\lambda 5411$ main component					
5 Jan 2013	6298.49765	401 ± 49	5.3 ± 2.4	17.6	−0.88 ± 0.71
2 Feb 2013	6326.30355	228 ± 59	7.8 ± 3.0	26.4	−1.2 ± 1.0
17 Feb 2013	6341.38141	321 ± 60	4.7 ± 2.8	15.6	−0.83 ± 0.91
22 Sep 2013	6557.63127	233 ± 78	5.5 ± 3.6	18.5	−1.1 ± 1.3
He II $\lambda 5411$ wing					
5 Jan 2013	6298.49765	−790 ± 330	36 ± 11	115.3	−2.8 ± 4.6
2 Feb 2013	6326.30355	−1200 ± 920	46 ± 33	144.5	−1.5 ± 5.6
17 Feb 2013	6341.38141	−760 ± 960	47 ± 35	150.4	−1.7 ± 8.5
22 Sep 2013	6557.63127	−100 ± 1100	93 ± 57	295.7	−6 ± 20

Table 5.8: CAL 83 highly ionized emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
[Fe x] $\lambda 6375$					
5 Jan 2013	6298.49765	289 ± 63	10.0 ± 3.2	33.2	−1.5 ± 1.3
2 Feb 2013	6326.30355	251 ± 46	7.9 ± 2.3	26.4	−1.5 ± 1.0
17 Feb 2013	6341.38141	248 ± 47	4.8 ± 2.3	15.6	−0.87 ± 0.91
O vi $\lambda 5290$					
26 Nov 2011	5891.56587	−200 ± 160	11.6 ± 6.8	37.7	−3.4 ± 8.0
11 Sep 2012	6181.60840	150 ± 130	7.8 ± 5.4	25.8	−2.0 ± 3.2
21 Mar 2013 (1)	6373.28286	177 ± 97	14.4 ± 4.0	47.1	−1.7 ± 2.7

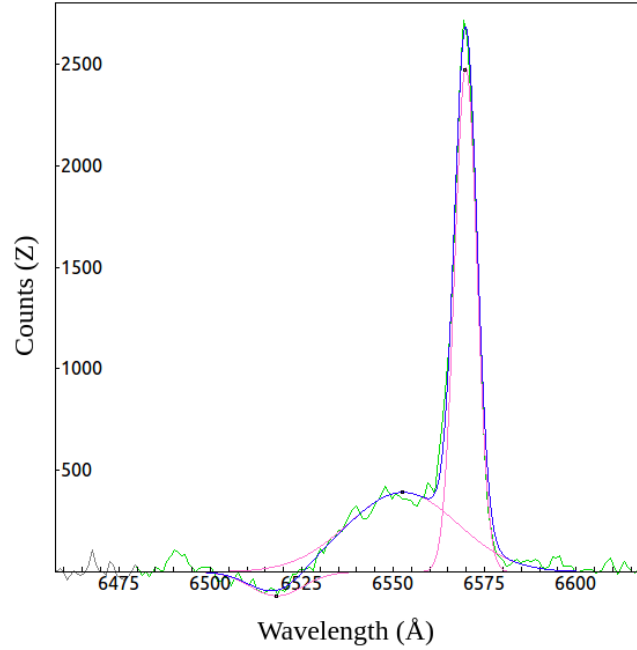


Fig. 5.11: Model fit to the CAL 83 H α spectral line observed on 5 January 2013. A combination of three Gaussian profiles was used: one for the main component, one for the broad emission wing and one for the blueshifted absorption. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

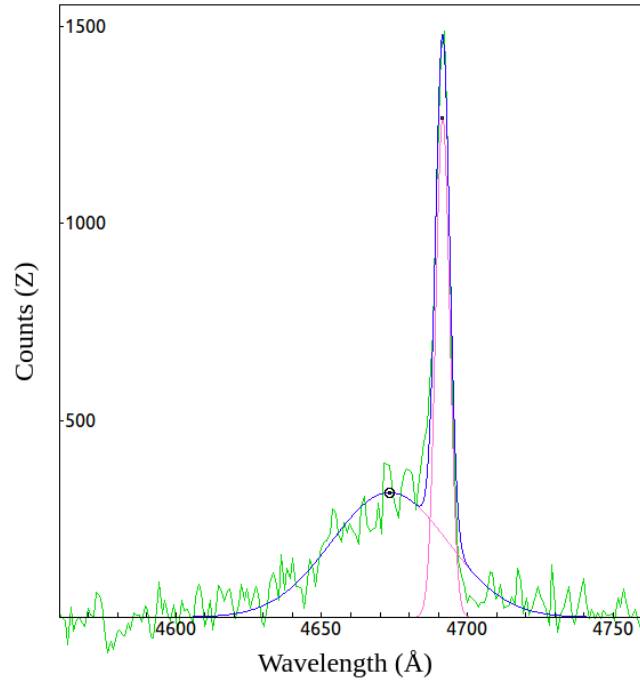


Fig. 5.12: Model fit to the CAL 83 He II $\lambda 4686$ emission observed on 5 January 2013. Two Gaussian profiles are used: one for the main component, and one for the combination of the He II $\lambda 4686$ wing and the blend of ionized carbon and nitrogen emission. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

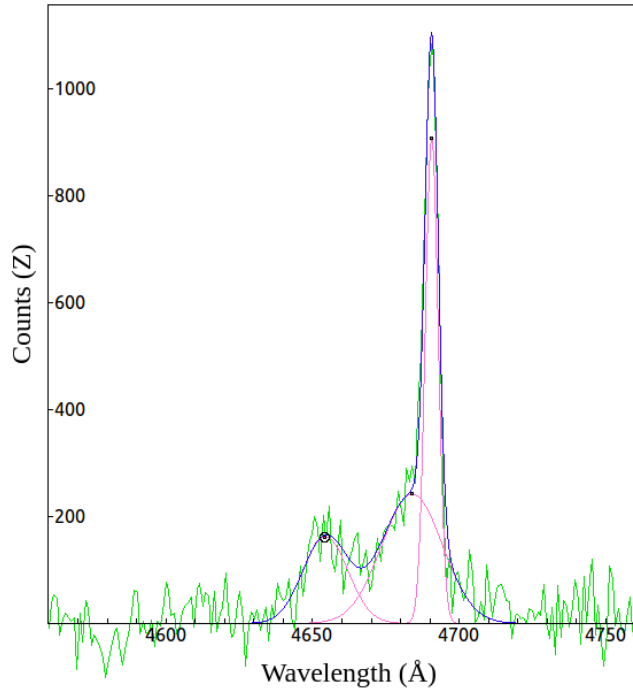


Fig. 5.13: Model fit to the CAL 83 He II $\lambda 4686$ emission observed on 17 February 2013. Three Gaussian profiles are used: one for the main component, one for the He II $\lambda 4686$ wing and one for the blend of ionized carbon and nitrogen emission. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

Table 5.9: Mean line ratios of the features in the spectra of CAL 83, calculated from the absolute values of their respective equivalent widths, and averaged over all the SALT observations. The equivalent width of the He II $\lambda 4686$ wing includes the contribution of the C III, C IV, N III blend.

Features compared	Mean line ratio
H α wing / H α main	0.96 $\pm$ 0.51
H α absorption / H α main	0.11 $\pm$ 0.1
H β main / H α main	0.221 $\pm$ 0.076
H γ main / H α main	0.13 $\pm$ 0.098
He II $\lambda 4686$ main / H α main	0.62 $\pm$ 0.33
He II $\lambda 5411$ main / H α main	0.097 $\pm$ 0.049
[Fe x] $\lambda 6375$ / H α main	0.122 $\pm$ 0.062
O VI $\lambda 5290$ / H α main	0.8 $\pm$ 1.7
H β wing / H β main	2.0 $\pm$ 2.1
H β absorption / H β main	1.5 $\pm$ 1.1
H γ main / H β main	0.88 $\pm$ 0.73
He II $\lambda 4686$ main / H β main	2.48 $\pm$ 0.70
He II $\lambda 5411$ main / H β main	0.56 $\pm$ 0.31
[Fe x] $\lambda 6375$ / H β main	0.82 $\pm$ 0.47
He II $\lambda 4686$ wing / He II $\lambda 4686$ main	3.2 $\pm$ 3.1
He II $\lambda 5411$ main / He II $\lambda 4686$ main	0.22 $\pm$ 0.11
[Fe x] $\lambda 6375$ / He II $\lambda 4686$ main	0.28 $\pm$ 0.15
O VI $\lambda 5290$ / He II $\lambda 4686$ main	0.77 $\pm$ 0.88
He II $\lambda 5411$ wing / He II $\lambda 5411$ main	2.9 $\pm$ 5.7
[Fe x] $\lambda 6375$ / He II $\lambda 5411$ main	1.30 $\pm$ 0.97

ratio over all the CAL 83 observations was calculated, and these are presented in Table 5.9. Note that all the ratios are positive for both the emission and absorption features, because the absolute values of W_λ were used.

In order to evaluate the evolution of the spectral properties with the overall optical flux, the W_λ and FWHM values from Tables 5.6, 5.7 and 5.8 were plotted together with a portion of the OGLE-IV lightcurve in Fig. 5.14 and Fig. 5.15 respectively. More information regarding the OGLE lightcurves will be provided in §6.2.

Measurements of the H β wing, as well as the H α and H β absorption features, were only obtained on 5 Jan and 2 Feb 2013. No variability within the 99.73% error margins was observed in the W_λ and FWHM values of these features between these two observations, and they are therefore not included in Fig. 5.14 and 5.15. Also, no variability was observed in the equivalent widths of the H α wing, the He II $\lambda 5411$ wing or the He II $\lambda 4686$ wing blended with C III, C IV and N III emission, and they are thus not shown in Fig. 5.14. Because the He II $\lambda 4686$ wing is blended with the C III, C IV and N III emission at ~ 4660 Å, its FWHM is not included in Fig. 5.15, as it is not really physically meaningful.

As can be seen in Fig. 5.14 and 5.15, the six SALT observations from 11 Sep 2012 to 21 Mar 2013 span one full optical high state according to the OGLE lightcurve.

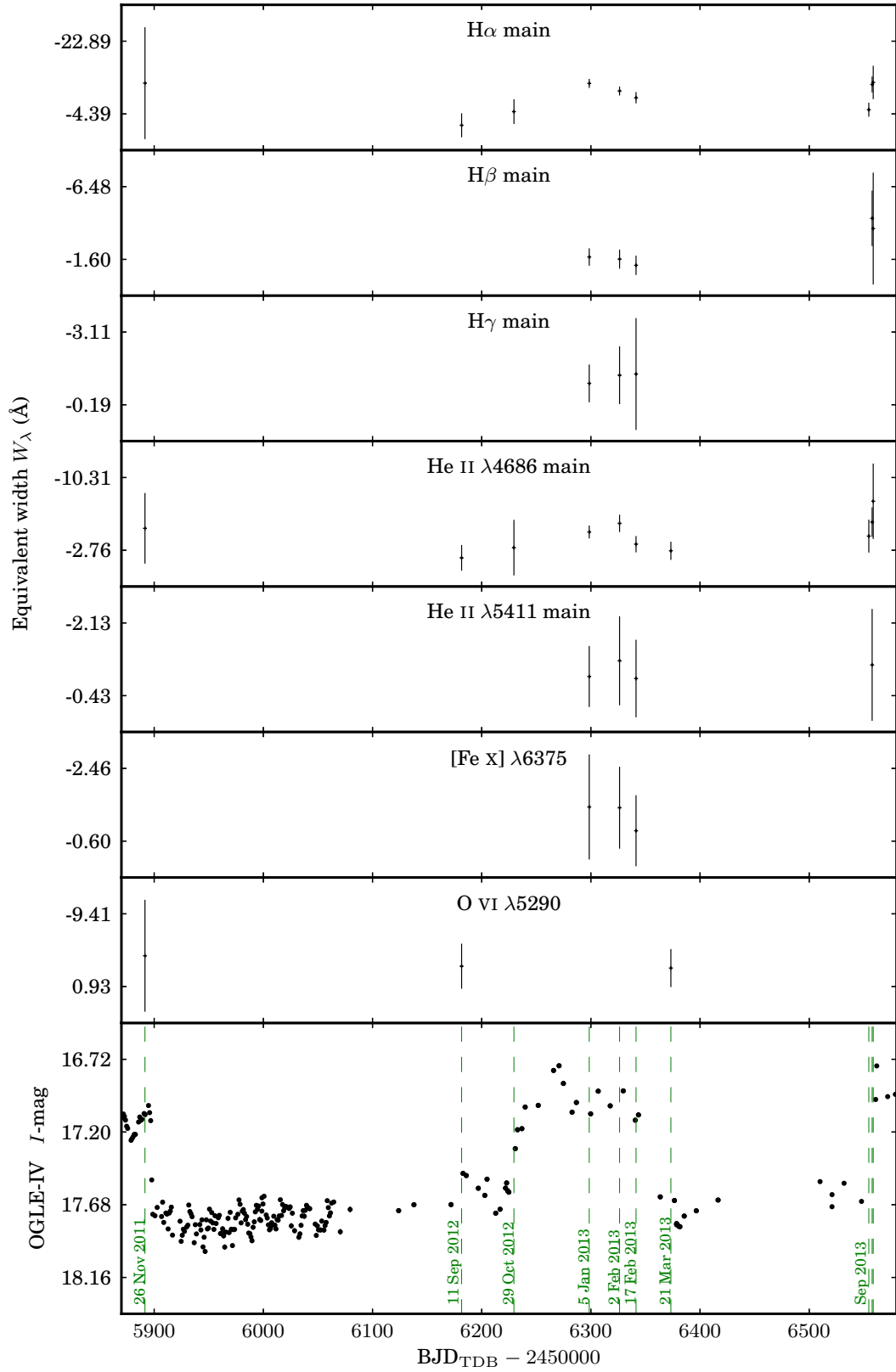


Fig. 5.14: The equivalent widths of spectral features of CAL 83 measured with SALT, as listed in Tables 5.6, 5.7 and 5.8, plotted together with the OGLE-IV lightcurve. The W_λ errors represent 99.73% confidence levels obtained by fitting Gaussian profiles. The OGLE-IV errors were also plotted, although they are mostly smaller than the data points.

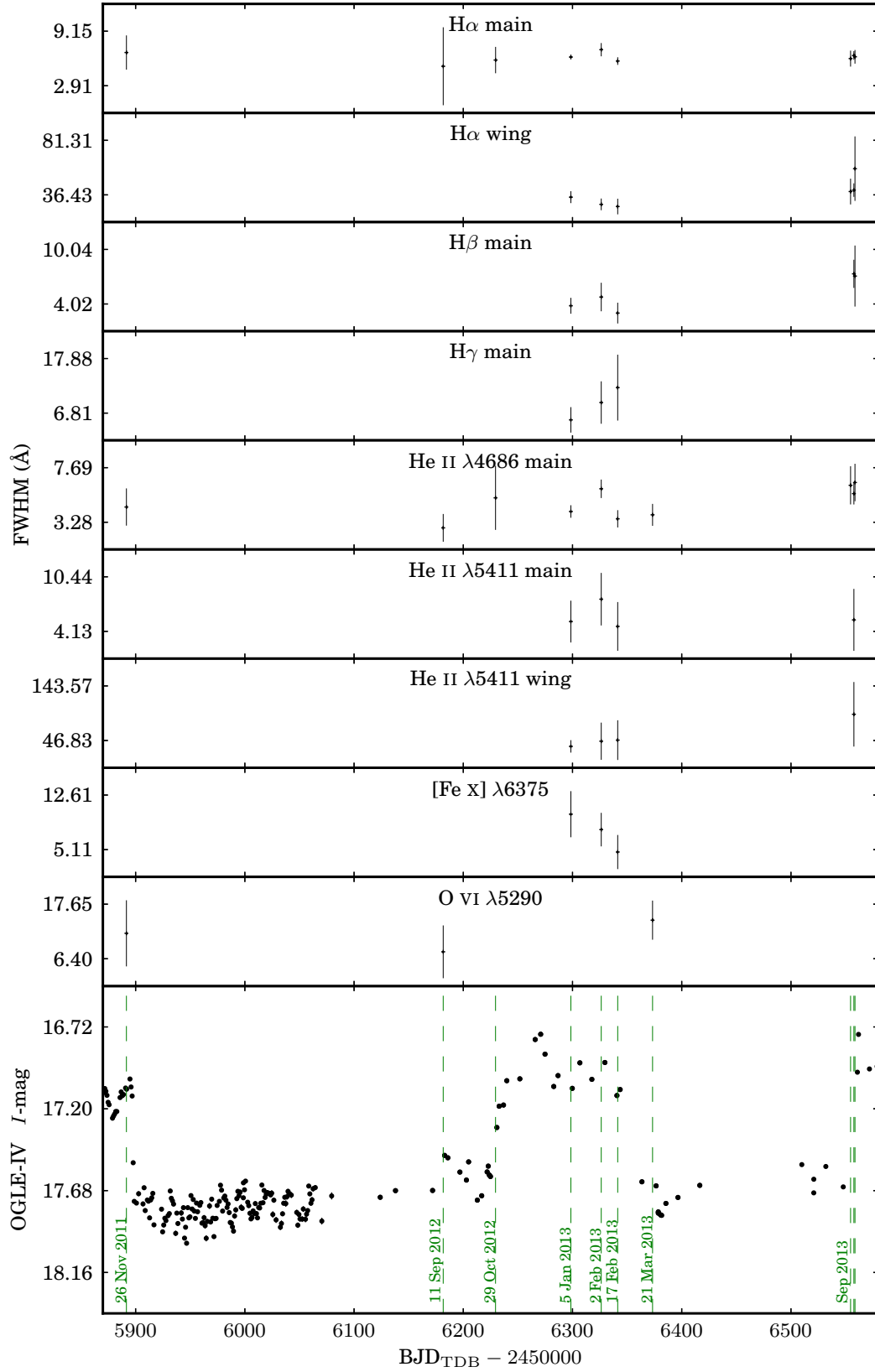


Fig. 5.15: The FWHM values of the CAL 83 spectral features measured with SALT, as listed in Tables 5.6, 5.7 and 5.8, plotted together with the OGLE-IV lightcurve. The FWHM errors represent 99.73% confidence levels obtained by fitting Gaussian profiles. The OGLE-IV errors were also plotted, although they are mostly smaller than the data points.

On 11 Sep 2012, the source was in an optical low state, while the 29 Oct 2012 observation occurred during the transition to an optical high state. The next three observations in Jan and Feb 2013 were performed towards the end of the optical high state. On 21 Mar 2013, the source was in an optical low state again.

From 11 Sep 2012 to 5 Jan 2013, there was a steady increase in W_λ of the H α main emission component as the optical brightness increased, and also a marginally significant decrease in W_λ from 5 Jan to 17 Feb 2013. A similar decrease may be present in the H β main component, but it is not statistically significant. The apparent increase in W_λ of the H γ main component is associated with large error bars. The He II $\lambda 4686$ main emission component showed a non-significant increase from 5 Jan to 2 Feb, and then a decrease towards 17 Feb and 21 Mar as the source entered the optical low state.

No OGLE measurements were obtained simultaneously with the Sep 2013 SALT observations, but from the interpolation between surrounding points it is assumed that the optical brightness was increasing from the optical low to the optical high state at that time. On 19 Sep 2013, W_λ of the H α main emission was significantly lower than in Jan 2013, which was during an optical high state. The equivalent widths of the H α and He II $\lambda 4686$ main emission lines seemed to increase from 19 to 23 Sep 2013, although this result is only marginally significant for the former and statistically insignificant for the latter.

Unfortunately, the S/N of the 26 Nov 2011 spectra (at the end of the previous optical high state) was too low to allow us to draw any definitive conclusions regarding a comparison of these W_λ measurements with the later measurements. A comparison of only the values themselves seems to indicate that the equivalent widths of H α and He II $\lambda 4686$ on 26 Nov 2011 were comparable to the values during the other optical high state in Jan 2013, but the error bars are very large.

According to the OGLE lightcurve, the optical brightness of the system increased slightly from 5 Jan to 2 Feb, then decreased on 17 Feb to a brightness slightly fainter than that on 5 Jan. In Fig. 5.15, it is interesting to note that the FWHM of the He II $\lambda 4686$ main component followed the same pattern. The He II $\lambda 4686$ FWHM was also significantly lower on 11 Sep 2012 during the optical low state, than on 2 Feb 2013 during the optical high state. A similar pattern can be observed at a marginal level in the FWHM of the H α main component, and also in the H β and He II $\lambda 5411$ main components, although the variability in the latter two lines is not significant over these 3 observations.

The FWHM of the H α wing is slightly higher on 22 Sep 2013, at the onset of the

next optical high state, than it was on 17 Feb 2013, at the end of the previous on-state. The same can be said for the FWHM of the H β main component. The He II λ 4686 main component FWHM is also markedly larger on 23 Sep 2013 than on 17 Feb 2013.

5.6.2 Discussion

Emission line production

The SALT spectra of CAL 83 have properties very similar to those reported in the literature since the first spectroscopic studies of the source in 1982 (refer back to §3.1.3). The strongest emission features are the hydrogen Balmer and He II emission lines, with weaker emission from ionized C, N and O. Note that the spectral analysis reported here focussed on the spectral properties of the central source only, as the background emission, including emission from the surrounding ionized nebula, has been subtracted.

The optical spectra are characterized by emission lines, which originate in optically thin regions in the source. In accreting compact binaries, the emission lines usually originate in the atmosphere/corona of the accretion disc, or in a wind or jet from the system. The H I emission will typically be associated with regions that have a relatively low temperature, i.e. below 157700 K (usually of the order of 15000 K) as hydrogen will be ionized above this temperature. As shown in Table 5.5, the ions with higher ionization states will be associated with higher temperature regions. This usually implies regions closer to the central X-ray source, although heating by shock fronts elsewhere in the system should also be considered.

The [Fe x] λ 6375 emission line is observed in many of the spectra. The presence of this line has for quite some time been considered as a signature of a supersoft phase in novae, caused by the photoionization of the surrounding corona by a hot ($T \gtrsim 4 \times 10^5$ K) central source (e.g. Krautter and Williams, 1989; Schwarz et al., 2011). Similarly, in SSSs, this line provides additional evidence of the presence of a hot, surface nuclear burning WD. The effective X-ray temperature of 550000 K derived by Lanz et al. (2005) would indeed be able to excite this line.

Before discussing the results of the optical spectra any further, the binary parameters of CAL 83 will first be investigated. Obtaining rough estimates of the source parameters, especially the location and dimension of the emission regions, will enable a better understanding of the spectral line characteristics.

Binary parameters of CAL 83

It has been shown in §3.1.3 that there is still considerable uncertainty in the values of the secondary mass and inclination of CAL 83. As discussed in §3.1.3, the radial velocity semi-amplitude of the He II $\lambda 4686$ line has been determined to be approximately 30 km s^{-1} . Assuming that this represents the WD motion, and adopting $P_{\text{orb}} = 1.047529 \text{ d}$ yields a very small mass function of $\sim 0.003 M_{\odot}$, when assuming an inclination of $i \sim 25^{\circ}$ (see Eq. (2.5) in §2.2). This corresponds to a secondary mass of $M_2 \sim 0.5 M_{\odot}$.

This mass is a full factor of 3 smaller than the estimate of $1.5\text{--}2 M_{\odot}$ of Van den Heuvel et al. (1992) for the donor mass in CAL 83. According to Eq. (2.19) in §2.3.2, the mass ratio $q = M_2/M_{\text{wd}}$ needs to be larger than 0.79 for fully conservative mass transfer to yield a shrinking Roche lobe, which is a requirement for the mass transfer expected to drive the nuclear burning on the WD surface. However, as discussed in the rest of that section, the loss of mass and angular momentum from the system can also contribute to the shrinking of the Roche lobe.

Although the spectral line profiles indicate that mass loss does occur in CAL 83, a crude estimate for the lower limit of the secondary mass can be obtained by considering the conservative case. With a WD mass of $1.3 M_{\odot}$, and $q > 0.79$, this yields $M_2 > 1.0 M_{\odot}$. It has been argued that, instead of using the radial velocity semi-amplitude of the He II $\lambda 4686$ line, it would be better to use higher ionization lines like the O VI lines to determine the true orbital motion of the WD, as they are expected to originate closer to the WD. However, Schmidtke et al. (2004) determined the radial velocity semi-amplitude of the O VI $\lambda 1032$ line to be comparable to that of the He II $\lambda 4686$ line (see Fig. 3.16). Even though there is still no guarantee that the O VI line yields a truly accurate representation of the orbital motion of the WD, there is thus no solid observational evidence that K_1 for the WD in CAL 83 is much larger than 30 km s^{-1} .

If $K_1 \sim 30 \text{ km s}^{-1}$ is adopted, setting $M_2 > 1.0 M_{\odot}$ yields an upper limit to the inclination of $i \lesssim 14^{\circ}$. A low inclination of $\sim 14^{\circ}$ would still be able to produce an orbital modulation in the optical due to the changing aspect of the irradiated accretion disc and donor, and, in particular, the presumed accretion disc rim (Schandl et al., 1997; Meyer-Hofmeister et al., 1997). Whether this inclination could exactly reproduce the optical lightcurves remains to be investigated by detailed modelling of the modulation of the emission from the rim. However, it is noted that there might be a narrow margin of secondary masses and associated inclination angles for which $K_1 \sim 30 \text{ km s}^{-1}$, $M_2 > 0.79 M_{\text{wd}} = 1.0 M_{\odot}$, and the inclination is not too low for the orbital modulation to be observed. The mass and angular momentum loss

associated with outflows will also contribute to driving the thermal time-scale mass transfer, as described by Eq. (2.18).

Adopting the lower limit of $M_2 \sim 1.0 M_\odot$, the orbital separation can be approximated from Eq. (2.1) as

$$a \sim 4.0 \times 10^{11} \left[\frac{M_1 + M_2}{(1.3 + 1.0) M_\odot} \right]^{1/3} \left(\frac{P_{\text{orb}}}{1.047529 \text{ d}} \right)^{2/3} \text{ cm} , \quad (5.3)$$

yielding Roche lobe radii of $R_{\text{L},1} \sim 1.6 \times 10^{11} \text{ cm}$ and $R_{\text{L},2} \sim 1.4 \times 10^{11} \text{ cm}$ when utilizing Eq. (2.7). A main sequence star with a mass of $1 M_\odot$ would have a radius of $\sim 7 \times 10^{10} \text{ cm}$, which is significantly smaller than the calculated Roche lobe radius of the secondary. For Roche lobe overflow to occur, a $1 M_\odot$ donor would therefore need to be significantly evolved. The envelope of such a secondary star would probably be convective, and the removal of material from such a convective envelope also causes an expansion, which can drive the Roche lobe overflow (e.g. Meintjes, 2002 and references therein).

The circularization radius corresponding to the binary parameters discussed above is $R_{\text{circ}} \sim 5.4 \times 10^{10} \text{ cm}$, according to Eq. (2.31). Assuming tidal truncation of the disc at $0.7R_{\text{L},1}$ (Papaloizou and Pringle, 1977), the outer disc limit is at $R_{\text{out}} \sim 1.1 \times 10^{11} \text{ cm}$. If the WD has a weak magnetic field of $\sim 10^5 \text{ G}$ (see §4.7.4), the disc can extend to the surface layers of the WD, and the inner disc radius can be approximated with $R_{\text{in}} \sim 10^9 \text{ cm}$.

Broadening of the main emission line components

The widths of the spectral lines obtained with the profile fitting should be considered in terms of the line broadening mechanisms discussed in §C.4. The instrumental broadening has already been removed from the presented FWHM and FWZI measurements by using Eq. (C.9). The thermal broadening FWHM for each spectral line was calculated with Eq. (C.11), and these values are also provided in Table 5.5. These values can be considered as rough upper limits for the amount of thermal broadening that can be expected in the respective lines, as they were calculated by using the temperature at which the relevant atom or ion will be ionized to the next ionization state. In most cases, the actual values would probably be much lower. E.g. for the $\text{H}\alpha$ line typically associated with emission regions at $\sim 15000 \text{ K}$, the thermal broadening FWHM is $\sim 0.7 \text{ \AA}$.

According to the tables of Vidal et al. (1973), the linear Stark broadening FWHM in material with a number density of $\sim 10^{15} \text{ cm}^{-3}$ is $\sim 0.1 \text{ \AA}$ for $\text{H}\alpha$ and $\sim 2 \text{ \AA}$ for $\text{H}\beta$. The Stark effect is known to be smaller for heavier atoms, and also smaller

for ions than for the neutral atoms of a certain element. Therefore, one can assume that the pressure broadening in He II and higher ionization lines will be less than in the hydrogen Balmer lines. It must be kept in mind that turbulent broadening could also play a role.

Considering the FWHM values of the strongest emission lines in Tables 5.6 and 5.7, i.e. H α and He II λ 4686, it is obvious that the widths of these lines can not be explained only by thermal and pressure broadening, and that another broadening mechanism dominates. Usually, in emission lines from accretion discs, rotational Doppler broadening is the dominant line broadening mechanism. Neglecting the other minor broadening effects and considering the line widths to be completely due to rotational broadening, therefore provides one with an upper limit to true rotational broadening associated with the line.

The inner and outer radii of the accretion disc in CAL 83 have been estimated in the previous section. According to Eq. (2.33), the circular velocities for material in Keplerian orbits at radii ranging from R_{out} to R_{in} , range from ~ 400 to ~ 4000 km s $^{-1}$. If the accretion disc was viewed edge-on (inclination angle $i \sim 90^\circ$), the wings of an observed emission line would therefore extend to ± 4000 km s $^{-1}$ from the line centre, corresponding to the maximum Doppler shift due to some particles at R_{in} moving directly towards the observer, and some directly away from the observer at any particular time. Therefore, the full width at zero intensity (FWZI) of the emission line would be ~ 8000 km s $^{-1}$. The emission line would also probably have a double-peaked structure (refer back to §2.4.3).

However, at smaller inclination angles, the radial velocity will only be a fraction of the above. Adopting $i \sim 14^\circ$, the FWZI projected along the line of sight would become $8000 \text{ km s}^{-1} \sin 14^\circ \sim 2000 \text{ km s}^{-1}$. For spectral lines emitted from regions extending to the inner disc around the WD, the expected FWZI is therefore of the order of $\sim 2000 \text{ km s}^{-1}$. The high temperatures in the inner regions of the disc would ionize the hydrogen, and the smallest radius associated with H I emission would therefore be somewhat further out, leading to a slightly smaller predicted FWZI.

The mean FWZI values of the main components of H α and He II λ 4686 are ~ 940 and ~ 1000 km s $^{-1}$ respectively. These values are of the same order of magnitude as the expected value quoted above. The fact that the measured values are slightly lower may simply indicate that the emission is concentrated slightly further out in the disc, or maybe that the inclination is even lower than $i \sim 14^\circ$. This crude approximation assumes that the line emission is completely derived from a perfectly symmetrical Keplerian accretion disc around the WD. However, this may not exactly be true.

For example, it has been suggested that the He II $\lambda 4686$ line originates in a region of interaction between a wind from the accretion disc and a wind from the heated side of the donor (e.g. Van den Heuvel et al., 1992). More detailed modelling of the accretion disc structure would be required to derive more rigid constraints from the line parameters.

Line emission from a wind

As explained in §2.8, two classes of outflows are usually distinguished in accreting binaries. The term “jet” usually refers to a narrow, well-collimated outflow from the compact object, approximately perpendicular to the orbital plane. The terminal velocity of the jet is approximately equal to the escape velocity v_{esc} of the compact object. The term “wind” refers to an outflow that is less collimated, and exhibits velocities ranging from approximately $0.1v_{\text{esc}}$ to v_{esc} . According to Eq. (2.98), the escape velocity from a region close to a $1.3 M_{\odot}$ WD is $v_{\text{esc}} \sim 5900 \text{ km s}^{-1}$. Considering a radiating mass element projected from the WD at a velocity equal to v_{esc} perpendicular to the orbital plane, its velocity component along the line of sight will be $v_{\text{esc}} \cos i \sim 5700 \text{ km s}^{-1}$, assuming $i \sim 14^{\circ}$ for CAL 83.

In addition to the “narrow” main components, the Balmer and He II emission lines of CAL 83 exhibit a very broad blue wing structure, as well as a P Cyg absorption feature blueshifted by several thousand kilometres per second. Because H α is the brightest line in the spectrum, and its wing is not “contaminated” by ionized C and N emission like the wing of He II $\lambda 4686$, this discussion will focus on the structure of the H α line. The equivalent width of the broad wing of H α is approximately equal to the equivalent width of the H α main component (Table 5.9). The equivalent width of the H α absorption feature has a value of at most 20% of the equivalent width of the main line.

In §2.8.3, it was explained how a spherically symmetric stellar wind can produce the characteristic P Cyg line profile. The component of the outflow producing the largest blueshift is the material moving from the central continuum emission source towards the observer. This component produces the highly blueshifted absorption feature. The emission component, centred on the rest wavelength, is observed from the regions where the line of sight of the observer do not intersect the central continuum source (see Fig. 2.18). If the central wavelength of the emission component is blueshifted relative to the rest wavelength, the outflow is not expected to be spherically symmetric, but rather directed in the approximate direction of the observer.

The H α wing was always blueshifted with respect to the H α main component, except on 23 Sep 2013, although negative values are included in the error margins of v_{SSB}

during the latter observation (see Table 5.6). It has been suggested by Crampton et al. (1987) that the variability in the shift of the wing structure with respect to the main component may be ascribed to a precession of the accretion disc with a period of ~ 69 d. Regular spectroscopic monitoring of CAL 83 is needed to better constrain the possible precession period.

The blueshifted absorption components of $H\alpha$ were only fitted for the observations on 5 Jan and 2 Feb 2013. For these observations, the mean velocity shift of the centre of the absorption component with respect to the centre of the main component is $-2400 \pm 110 \text{ km s}^{-1}$. The mean FWZI measured for the absorption feature is $\sim 3800 \text{ km s}^{-1}$, therefore its blue wing blends into the continuum at a velocity of $(-2400 - 0.5 \times 3800) \text{ km s}^{-1} \sim -4300 \text{ km s}^{-1}$. Comparing this value with the projected value of v_{esc} shows that the upper limit to the terminal velocity of the wind is $\sim 0.8v_{\text{esc}}$.

The mean FWHM of the $H\alpha$ wing is $1730 \pm 240 \text{ km s}^{-1}$. If the velocity distribution in the wind is uniform, the large spread in velocities indicate a wind with a large opening angle, i.e. containing a large range of angles relative to the line of sight, yielding a large range of velocities projected on the line of sight. A substantial gradient in intrinsic velocities within the wind could also produce the highly widened structure for a reasonably well-collimated jet. The terminal velocity of $\sim 0.8v_{\text{esc}}$ also allows the outflow to be classified as either a “wind” or a “jet” according to the differentiation above.

However, the detection of both a well-collimated jet and a less collimated outflow in the similar source RXJ0513 (see §5.7.1), suggests that a well-collimated outflow in CAL 83, if present, might also have presented well-defined blue- and redshifted satellite features rather than only the observed broad outflow feature. As discussed in §4.7.2, there are indications that the accretion rate in CAL 83 may be lower compared to e.g. RXJ0513, most probably placing CAL 83 below the limiting condition of $\gtrsim 10^{-6} \text{ M}_{\odot}\text{yr}^{-1}$ imposed by Soker and Lasota (2004) as a prerequisite for the presence of collimated jets in disc-accreting WD binaries (refer back to §2.8). Therefore, the outflow in CAL 83 is most likely best described as a wind with a large opening angle.

The wind may either be a radiation driven wind, or a magnetocentrifugal outflow, as discussed in §2.8. A relatively weak WD magnetic field of $\sim 10^5 \text{ G}$ formed part of a plausible LIMA-type model for the $\sim 67 \text{ s}$ X-ray periodicity (§4.7.4). It is highly probable that a magnetocentrifugal wind can be present in the system as a natural consequence of MHD turbulence.

Other long-term trends

As shown in Fig. 5.14 and 5.15, the W_λ and FWHM values of several of the spectral features of CAL 83 exhibit some variability between the different observations. Time-resolved spectroscopic observations covering a significant fraction of the orbital period were not obtained, and the number of observations is not sufficient to draw solid conclusions regarding any orbital phase dependence when folding these quantities on the orbital ephemeris. However, significant orbital phase-dependent variability in W_λ and FWHM of the spectral lines have not been reported in the literature. Another possibility is that these changes may be related to the superorbital cycles in CAL 83.

A comparison of the equivalent widths in Fig. 5.14 with the long-term OGLE-IV lightcurve revealed a general tendency of the equivalent widths to change in step with the overall optical brightness of the system. This is to be expected, as the extended accretion disc and possibly additional winds blowing from the WD surface will enhance the intensity of the optically thin line emission from the system.

A similar, though less significant, trend was observed when comparing the FWHM values of the emission lines with the OGLE lightcurve. It is not trivial to explain the association of a wider accretion disc line with an extended envelope. If anything, the expanded envelope should obscure the inner regions of the disc associated with higher circular velocities. However, the enhanced accretion rate could be associated with enhanced winds from the system, with the onset of the so-called “accretion wind” according to the Hachisu and Kato (2003a) model. Line emission from these additional high-velocity components could therefore contribute to wider emission lines, both for the main components and wing structures.

Variability in the O VI $\lambda 5290$ emission

The W_λ values of the O VI $\lambda 5290$ line did not exhibit significant variability between the 3 observations where a reasonable fit could be obtained, and it is also noted that these values are statistically compatible with a non-detection (see Table 5.8). However, a qualitative evaluation of the observed intensity of this line by scrutinizing Fig. 5.9 revealed an interesting trend in the O VI emission.

On 26 Nov 2011, at the end of an optical high state, the intensity of the O VI $\lambda 5290$ emission was comparable to or even slightly higher than that of its neighbour, He II $\lambda 5411$. The O VI $\lambda\lambda 3811, 3834$ lines were also detected at a marginal level. On 11 Sep 2012, during the optical low state, the situation was similar. The situation on 29 Oct 2012 during the transition to the optical high state is unclear, as all 4 of these lines had almost disappeared into the noise level during the latter observation.

However, during the optical high state observations on 5 Jan, 2 Feb and 17 Feb 2013, where the S/N was much higher, O VI $\lambda 5290$ was barely discernible above the noise level at all. The He II $\lambda 5411$ line was relatively strong during these observations. Unfortunately, the O VI $\lambda\lambda 3811, 3834$ lines were not included in the spectral range during these 3 observations, so the behaviour of these lines during the optical high state is not clear.

On 21 Mar 2013, after the commencement of the next optical low state, the O VI $\lambda 5290$ line was back again at a level comparable to that of He II $\lambda 5411$. The O VI $\lambda\lambda 3811, 3834$ lines were also discernible during this observation. The transition to the next optical high state occurred in Sep 2013. The O VI $\lambda 5290$ line was still detectable on 19 Sep 2013, but at a slightly lower level than He II $\lambda 5411$. In the 22 Sep spectrum, the S/N was comparable to that of 19 Sep, but the O VI $\lambda 5290$ line was undetectable. It was also not observed on 23 Sep, although it is noted that the S/N of the latter spectrum was somewhat worse. The He II $\lambda 5411$ line was clearly still present at more or less the same level in the latter two observations.

The behaviour of the O VI $\lambda 5290$ line in the previously published spectroscopic observations summarized in Chapter 3 has also been investigated. Crampton et al. (1996) obtained an optical spectrum of CAL 83 at the CTIO 4-m telescope in November 1994 during their spectroscopic observations of RXJ0513. At this time ($\text{BJD}_{\text{TDB}} \sim 2449668$), CAL 83 was undergoing a transition from an intermediate optical state to an optical low state, according to the MACHO lightcurve that is presented in Fig. 6.6 of §6.2 of this thesis. As shown in Fig. 3.38, the O VI $\lambda 5290$ emission intensity is quite high at this time, i.e. at a level comparable to that of He II $\lambda 5411$.

Cowley et al. (1998) presented optical spectra of CAL 83 obtained with the CTIO 4-m telescope in early November 1996 (see Fig. 3.14). The BJD_{TDB} corresponding to these observations is ~ 2450392 , and a comparison with the MACHO lightcurve (Fig. 6.6) indicates that the source was in the transition phase from an optical low to an optical high state at the time. In the spectrum of Cowley et al. (1998), the O VI $\lambda 5290$ line is present at a slightly weaker level than the He II $\lambda 5411$ line.

The VLT observations of Gruyters et al. (2012) near the end of 2005 occurred during an optical low state according to the OGLE-III lightcurve in Fig. 6.6. As shown in Fig. 3.23, the intensity of O VI $\lambda 5290$ was comparable to the intensity of He II $\lambda 5411$ at the time.

As discussed before (e.g. §3.1), current observational evidence indicates that the X-

ray and optical flux from the source are anti-correlated. This has been explained by the expansion/contraction of the WD envelope. During the optical low state, the envelope is in a contracted state, and the peak of emission is in the supersoft X-ray band. A presumed increase in the accretion rate causes the WD envelope to expand by several factors, shifting the peak towards the ultraviolet, and causing an X-ray low, optical high state.

The O VI emission lines are expected to originate in the inner accretion disc, close to the WD. The O VI ion has a large ionization potential (see Table 5.5), and can thus be assumed to form in close proximity to the central, hot WD, where the photons forming part of the high energy tail of the thermal emission may ionize some of the oxygen to high ionization states. The general trend observed in the SALT spectra, and supported by published data, is that the O VI $\lambda 5290$ emission is present at a significant level during the optical low states, but is much weaker during the optical high states.

This phenomenon can be readily explained as part of the expansion/contraction model of the WD. When the envelope is in the optical low state with a contracted envelope, optically thin regions of the inner disc is visible, and the O VI $\lambda 5290$ line is observed. Upon entering the optical high state, the envelope expands, the O VI emission is largely absorbed by the expanded envelope surrounding it, and the O VI $\lambda 5290$ line becomes much weaker than before.

Within this context, one may ask why the O VI $\lambda 5290$ line was already detected at the end of the optical high state on 26 Nov 2011 before the optical flux had decreased. It is possible that the WD photosphere had at that stage already contracted enough to start revealing the O VI $\lambda 5290$ emission region, while the accretion disc was still in a brighter state, and still shrinking to its optical low state dimensions after the preceding episode of high-rate mass accretion.

This trend was not observed in the H I and He II emission, as these lines can be produced much further away from the WD, and are therefore not attenuated by the expanding envelope. The trend was also not detected from the ionized C and N emission. It is noted that the ionization temperatures required for the C III, C IV, N III and N V lines are still lower than for O VI, therefore the C and N lines may originate in slightly cooler regions of the disc somewhat further away from the WD surface, or perhaps in a wind excited from the secondary star.

Although the detection of possible cyclic variability in the O VI $\lambda 5290$ line intensity reported here is of a qualitative nature, we propose that this definitely warrants further investigation. Solid confirmation of this trend could be obtained by regular

spectroscopic monitoring at different phases of the superorbital cycle.

5.7 RX J0513.9-6951

5.7.1 Results

The calibrated, continuum-subtracted spectra of RX J0513.9-6951 (RXJ0513) obtained during the observations listed in Table 5.3 are shown in Fig. 5.16, with the detected emission lines marked on each spectrum. The properties of all the emission lines that have been detected previously in RXJ0513 are listed in Table 5.10 (see §3.2.3). For reference, their positions are also indicated in the spectra in Fig. 5.16, regardless of whether or not they were observed in the respective spectra. The line labels in a particular spectrum were offset by a fixed velocity equal to the mean velocity of the fitted lines in that particular spectrum.

The spectra of RXJ0513 are in many ways similar to those of CAL 83. The $H\alpha$ and He II $\lambda 4686$ emission lines are the most prominent features, and lines of the H I Balmer series up to H8, as well as the He II lines at 4026, 4200, 4541 and 5411 and 6683 Å were also detected. The $H\alpha$ and $H\beta$ lines are expected to be blended with He II $\lambda 6560$ and He II $\lambda 4859$ respectively.

In addition to “narrow” central components, the Balmer and He II lines exhibit broad bases/wings with a full width of $\sim 2000 \text{ km s}^{-1}$. In many of the spectra, in particular on 27 Aug 2012, these lines also exhibit blueshifted (S^-) and redshifted (S^+) satellite features at $\pm 4000 \text{ km s}^{-1}$. To illustrate these features, the $H\beta$, $H\alpha$ and He II $\lambda\lambda 4686, 5411$ lines observed on 27 August 2012 are plotted in terms of velocity relative to the central line wavelength in Fig. 5.17.

The S^- and S^+ features were the most prominent on 27 Aug 2012. On 29 Oct 2012, these features were less pronounced, but the slightly lower S/N also made the detection more difficult. However, the $H\alpha$ S^- feature now seemed to appear in absorption. On 18 Sep 2013, the S/N was comparable to that of 27 Aug 2012, but the satellite features were still weaker, with $H\alpha$ S^- possibly slightly absorbed. The highest S/N spectrum is the one on 22 Sep 2013, where there is virtually no trace of the satellite features in any of the lines. The S/N on 25 Sep 2013 was again comparable to that on 29 Oct 2012, and while the possible S^+ features can be discerned in emission, S^- in $H\alpha$ and perhaps also in $H\beta$ seem to be absorbed again.

Emission of highly ionized species are also present in the spectra, i.e. O VI $\lambda\lambda 3811, 3834$, faint traces of N V $\lambda\lambda 4602, 4620$, a combination of C III, C IV, N III and Si IV (blended with the blue wing of He II $\lambda 4686$), O VI $\lambda 5290$, O V $\lambda 5598$ and C IV $\lambda\lambda 5801, 5912$.

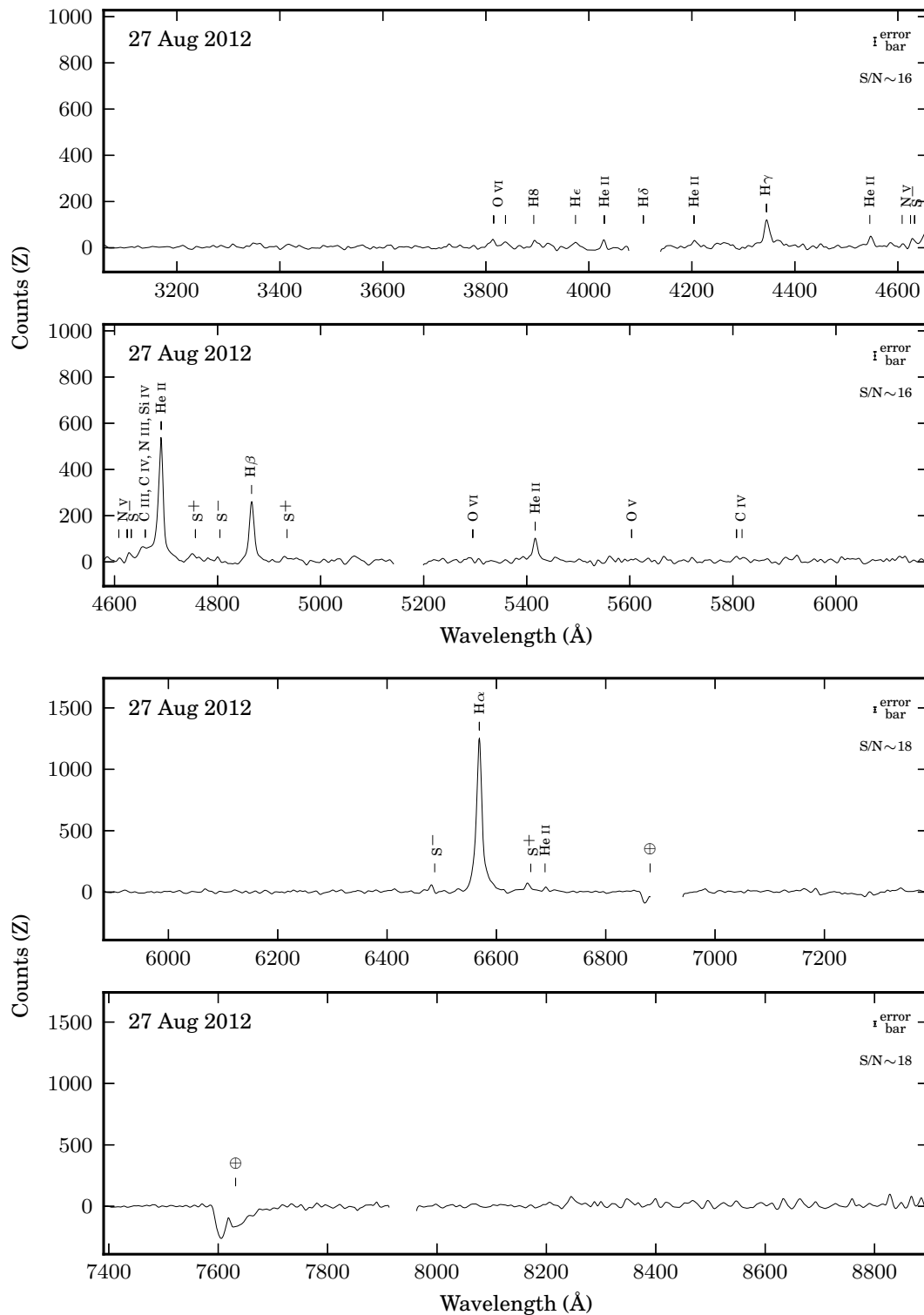


Fig. 5.16: SALT RSS spectra of RXJ0513. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

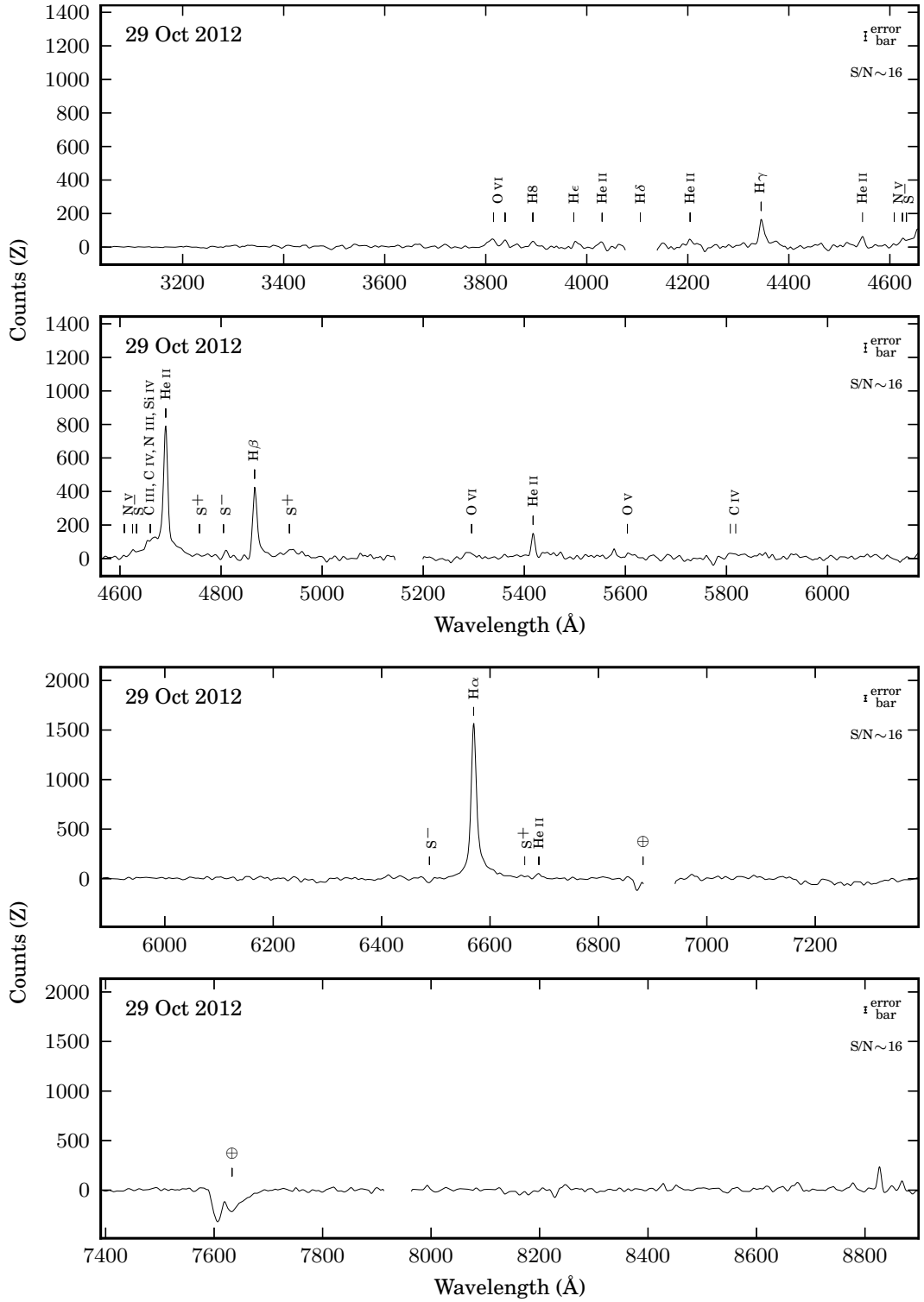


Fig. 5.16: (continued) SALT RSS spectra of RXJ0513. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

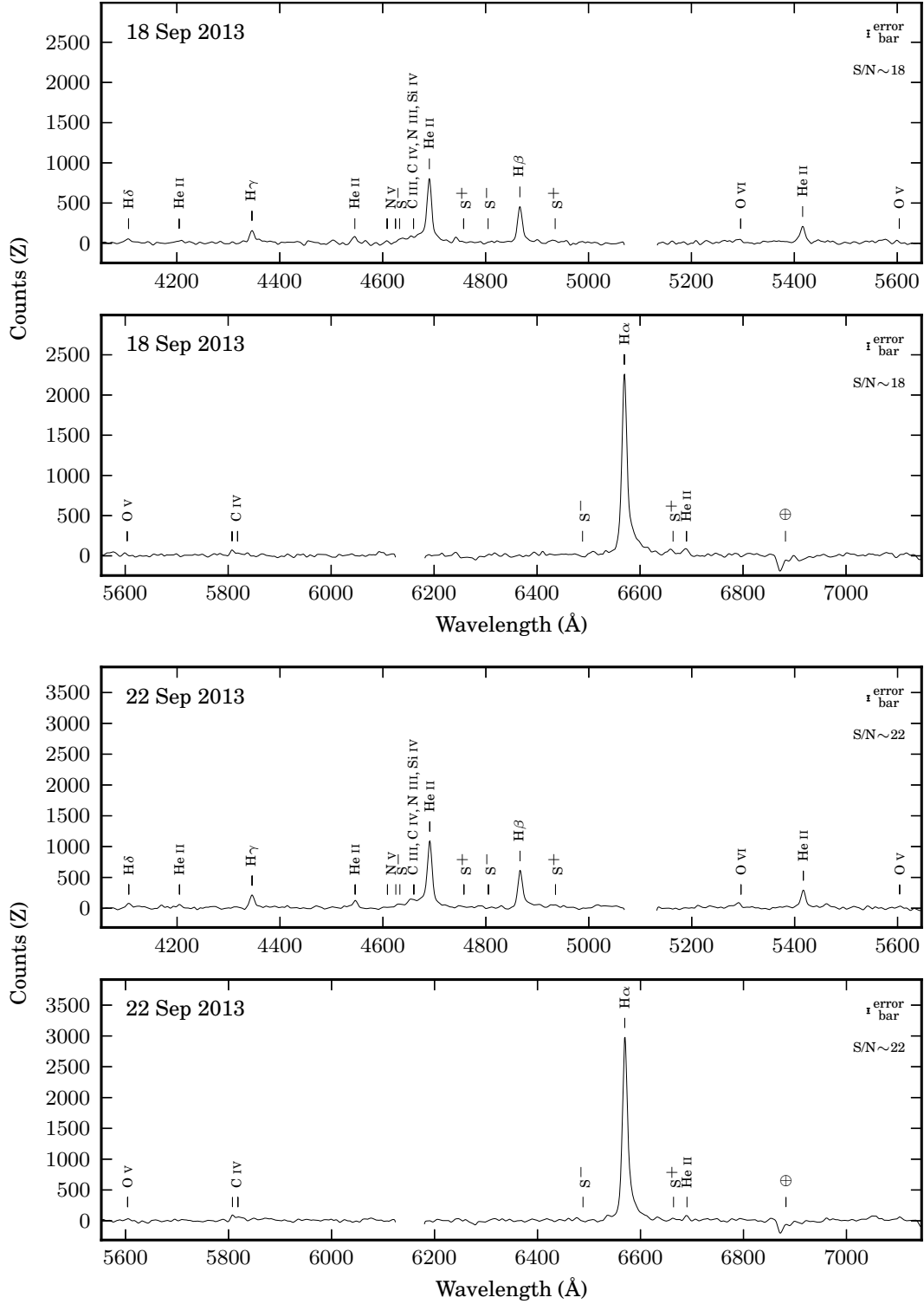


Fig. 5.16: (continued) SALT RSS spectra of RXJ0513. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

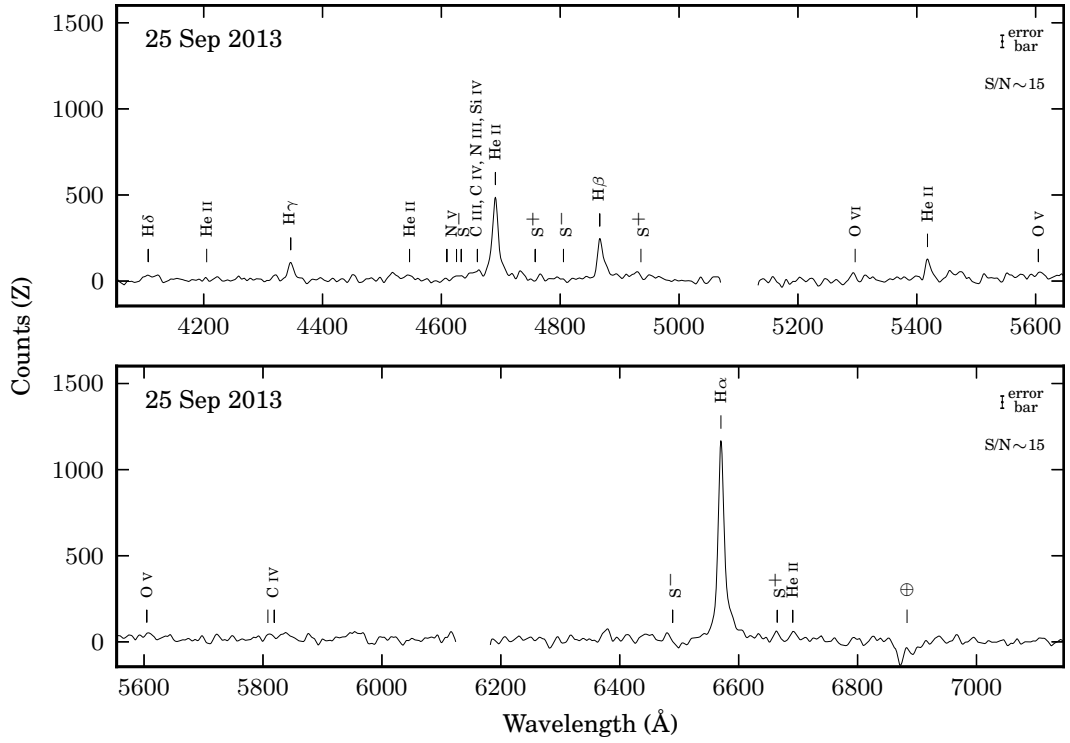


Fig. 5.16: (continued) SALT RSS spectra of RXJ0513. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

Table 5.10: Standard properties of the emission lines observed in RXJ0513. The third column indicates the temperature range within which the ion can exist in the relevant ionization state without being excited to an even higher state, according to the ionization potentials and assuming thermal ionization. The last column represents the thermal broadening FWHM according to Eq. (C.11) for the single lines, corresponding to the highest temperature in the third column. The rest wavelengths and ionization potentials were obtained from the NIST Atomic Spectra Database.

Element and ionization state	Rest wavelength λ_0 (Å)	Temperature range (K)	Thermal broadening FWHM (Å)
H I (H α) with bipolar satellite features (Blended with He II λ 6560)	6562.79	<157700	1.9
H I (H β) with blue absorption and bipolar satellite features (Blended with He II λ 4859)	4861.35	<157700	1.4
H I (H γ)	4340.472	<157700	1.2
H I (H δ) (Blended with Si IV $\lambda\lambda$ 4089 4116)	4101.734	<157700	1.2
H I (H ϵ)	3970.075	<157700	1.1
H I (H8)	3889.064	<157700	1.1
He II	4026	285200-631200	1.1
He II	4200	285200-631200	1.2
He II	4541	285200-631200	1.3
He II with blue wing and bipolar satellite features	4686	285200-631200	1.3
He II	5411	285200-631200	1.5
He II	6683	285200-631200	1.9
C III blend (?)	4070	282900-555500	
Blend of C III, C IV, N III, and Si IV emission	~4630-4680	282900-555500 555500-748100 343400-550400 388500-523600	
C IV	5801.33	555500-748100	1.0
C IV	5811.98	555500-748100	1.0
N V	4603.73	898700-1136000	0.9
N V	4619.98	898700-1136000	0.9
O V	5597.91	898000-1321000	1.1
O VI	3811.35	1321000-1602000	0.9
O VI	3834.24	1321000-1602000	0.9
O VI	5290	1321000-1602000	1.2

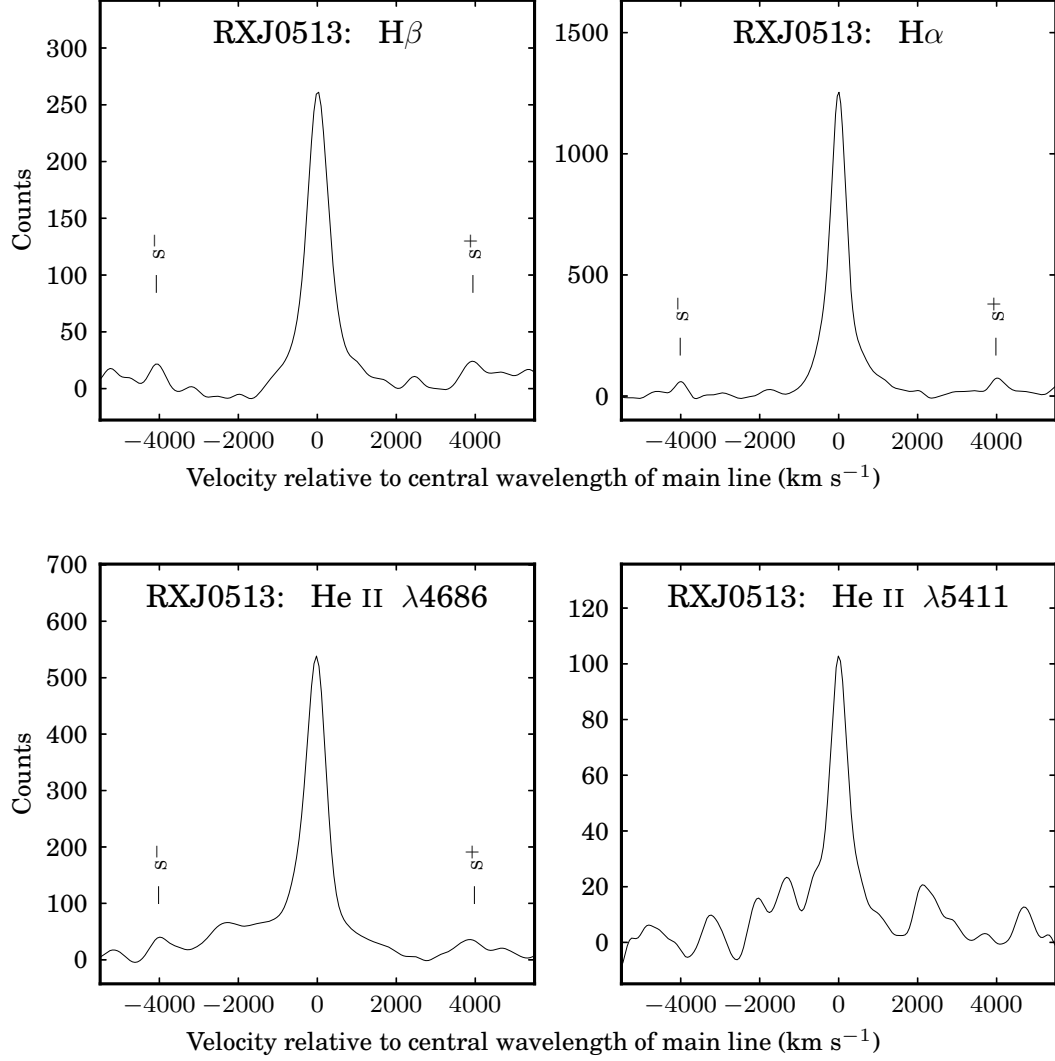


Fig. 5.17: The spectral regions around the H β , H α and He II $\lambda\lambda$ 4686,5411 emission lines in the RXJ0513 spectrum on 27 August 2012, plotted in terms of velocity relative to the central line position. Each line has a narrow core superimposed on a broad base, and satellite emission features are visible at ± 4000 km s⁻¹ in H β , H α and He II λ 4686. Emission of C III, C IV, N III and Si IV is expected to contribute to the blue wing of He II λ 4686.

Table 5.11: RXJ0513 H α emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing and satellite components it is with respect to the central wavelength of the relevant main component.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
Hα main component					
27 Aug 2012	6166.65683	267.8 $\pm$ 3.9	4.54 $\pm$ 0.26	15.2	−17.9 $\pm$ 2.2
29 Oct 2012	6229.60418	316.0 $\pm$ 5.8	5.07 $\pm$ 0.36	17.1	−21.1 $\pm$ 2.9
18 Sep 2013	6553.62610	301.6 $\pm$ 5.1	5.60 $\pm$ 0.31	18.6	−20.0 $\pm$ 2.6
22 Sep 2013	6557.56369	302.4 $\pm$ 4.5	5.39 $\pm$ 0.29	18.6	−20.8 $\pm$ 2.2
25 Sep 2013	6560.61358	334.2 $\pm$ 6.9	5.35 $\pm$ 0.43	17.6	−11.7 $\pm$ 2.4
Hα wing					
27 Aug 2012	6166.65683	−1 $\pm$ 30	24.3 $\pm$ 2.0	78.1	−27.2 $\pm$ 7.7
29 Oct 2012	6229.60418	145 $\pm$ 69	32.5 $\pm$ 3.9	102.8	−29 $\pm$ 11
18 Sep 2013	6553.62610	312 $\pm$ 85	39.0 $\pm$ 4.2	124.0	−23 $\pm$ 11
22 Sep 2013	6557.56369	217 $\pm$ 51	27.5 $\pm$ 2.4	87.8	−22.8 $\pm$ 6.9
25 Sep 2013	6560.61358	208 $\pm$ 89	35.7 $\pm$ 4.8	113.2	−15 $\pm$ 11
Hα blueshifted jet					
27 Aug 2012	6166.65683	−3990 $\pm$ 44	3.7 $\pm$ 2.3	13.3	−1.1 $\pm$ 1.1
Hα redshifted jet					
27 Aug 2012	6166.65683	4026 $\pm$ 69	7.6 $\pm$ 3.6	25.7	−2.0 $\pm$ 2.1

The [Fe x] λ 6375 line was not detected.

Detailed fitting as described in §5.5.4 and §5.5.5 was only performed for the strongest H I and He II lines in each spectrum. The results are presented in Tables 5.11, 5.12, 5.13 and 5.14.

The H α and H β spectral lines were fitted with two Gaussian profiles for all the observations: one for the main emission line and another for the broad emission wing. On 27 Aug 2012, the redshifted and blueshifted jet satellite features in H α and H β were clearly visible, and each feature was fitted with an additional Gaussian profile, as shown in Fig. 5.18 for H α . The H γ main emission component was fitted with a single Gaussian profile for all the observations, with an additional Gaussian component for the wing structure on 29 Oct 2012 and 18 Sep 2013.

For the spectra of 18 Sep 2013 and 25 Sep 2013, the He II λ 4686 emission was fitted with two Gaussian profiles: one for the main emission line and another for the combination of the He II λ 4686 wing structure and the blend of C III, C IV, N III and Si IV. For the spectra of 27 Aug 2012, 29 Oct 2012 and 22 Sep 2013, the He II λ 4686 wing and the C III, C IV, N III, Si IV blend were fitted with two separate Gaussian profiles. In addition, for the 27 Aug 2012 spectrum, the pronounced bipolar satellite

Table 5.12: RXJ0513 H β emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing and satellite components it is with respect to the central wavelength of the relevant main component.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
Hβ main component					
27 Aug 2012	6166.64912	284 ± 17	3.64 ± 0.86	12.8	−3.4 ± 1.3
29 Oct 2012	6229.59298	308 ± 19	5.23 ± 0.9	17.9	−6.3 ± 2.0
18 Sep 2013	6553.62610	321 ± 25	6.1 ± 1.1	20.5	−5.6 ± 1.9
22 Sep 2013	6557.56369	310 ± 18	5.53 ± 0.79	18.5	−5.8 ± 1.5
25 Sep 2013	6560.61358	316 ± 32	3.6 ± 1.4	11.7	−2.5 ± 1.2
Hβ wing					
27 Aug 2012	6166.64912	60 ± 110	18.2 ± 5.3	58.1	−5.5 ± 5.4
29 Oct 2012	6229.59298	580 ± 420	24 ± 12	76.4	−4.0 ± 7.1
18 Sep 2013	6553.62610	400 ± 1000	63 ± 48	200.1	−5 ± 16
22 Sep 2013	6557.56369	820 ± 410	39 ± 13	124.9	−5.5 ± 8.5
25 Sep 2013	6560.61358	520 ± 370	35 ± 12	110.3	−6 ± 11
Hβ blueshifted jet					
27 Aug 2012	6166.64912	−4080 ± 100	3.2 ± 4.0	10.8	−0.42 ± 0.96
Hβ redshifted jet					
27 Aug 2012	6166.64912	3940 ± 230	8.2 ± 8.8	26.6	−0.8 ± 2.3

Table 5.13: RXJ0513 H γ emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing component it is with respect to the central wavelength of the relevant main component.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
Hγ main component					
27 Aug 2012	6166.64912	301 ± 48	4.6 ± 1.8	14.8	−2.0 ± 1.4
29 Oct 2012	6229.59298	314 ± 35	3.5 ± 1.3	11.9	−1.9 ± 1.2
18 Sep 2013	6553.62610	376 ± 86	6.4 ± 3.3	21.5	−2.6 ± 1.8
22 Sep 2013	6557.56369	357 ± 60	7.3 ± 2.0	23.4	−3.7 ± 1.7
25 Sep 2013	6560.61358	410 ± 100	8.0 ± 3.5	26.4	−3.7 ± 2.7
Hγ wing					
27 Aug 2012	6166.64912	490 ± 540	46 ± 19	145.8	−6 ± 13
29 Oct 2012	6229.59298	600 ± 690	46 ± 24	144.8	−4 ± 13
18 Sep 2013	6553.62610	900 ± 1400	49 ± 45	157.2	−4 ± 12

Table 5.14: RXJ0513 He II emission: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the wing and satellite components it is with respect to the central wavelength of the relevant main component. In cases where the He II $\lambda 4686$ wing and C III, C IV, N III, Si IV blend were fitted separately, two measurements are given. The first is applicable to the wing, and the second to the carbon-nitrogen-silicon blend.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
He II $\lambda 4686$ main component					
27 Aug 2012	6166.64912	308.3 ± 8.7	2.68 ± 0.41	8.9	−5.3 ± 1.1
29 Oct 2012	6229.59298	310 ± 10	2.69 ± 0.63	9.9	−5.5 ± 1.2
18 Sep 2013	6553.62610	320 ± 15	5.37 ± 0.63	17.6	−8.9 ± 1.9
22 Sep 2013	6557.56369	334 ± 12	5.19 ± 0.54	17.6	−9.2 ± 1.6
25 Sep 2013	6560.61358	331 ± 24	5.06 ± 0.98	17.6	−7.0 ± 2.1
He II $\lambda 4686$ wing + C III, C IV, N III, Si IV blend					
27 Aug 2012	6166.64912	−155 ± 80	10.6 ± 2.5	34.5	−5.6 ± 3.3
	+	−820 ± 340	61 ± 13	195.1	−13 ± 18
29 Oct 2012	6229.59298	−57 ± 85	8.7 ± 4.4	27.8	−4.2 ± 2.7
	+	−590 ± 320	69 ± 15	217.2	−18 ± 21
18 Sep 2013	6553.62610	−630 ± 390	62 ± 16	197.2	−14 ± 16
22 Sep 2013	6557.56369	−220 ± 260	38 ± 12	120.1	−12.2 ± 8.5
	+	−2420 ± 260	17.5 ± 9.7	56.6	−3.0 ± 3.9
25 Sep 2013	6560.61358	−160 ± 510	64 ± 24	202.1	−14 ± 19
He II $\lambda 4686$ blueshifted jet					
27 Aug 2012	6166.64912	−4004 ± 80	2.6 ± 3.0	8.9	−0.46 ± 0.8
He II $\lambda 4686$ redshifted jet					
27 Aug 2012	6166.64912	3940 ± 330	16 ± 13	51.2	−1.5 ± 4.5
He II $\lambda 5411$ main component					
27 Aug 2012	6166.64912	298 ± 30	4.5 ± 1.4	15.8	−2.0 ± 1.5
29 Oct 2012	6229.59298	345 ± 27	4.0 ± 1.1	13.9	−2.2 ± 1.4
18 Sep 2013	6553.62610	283 ± 70	8.2 ± 3.0	27.3	−2.7 ± 2.3
22 Sep 2013	6557.56369	325 ± 40	6.3 ± 1.7	20.5	−2.7 ± 1.5
25 Sep 2013	6560.61358	388 ± 76	6.4 ± 3.2	21.5	−1.9 ± 2.1
He II $\lambda 5411$ wing					
27 Aug 2012	6166.64912	−170 ± 760	63 ± 39	201.0	−4 ± 18

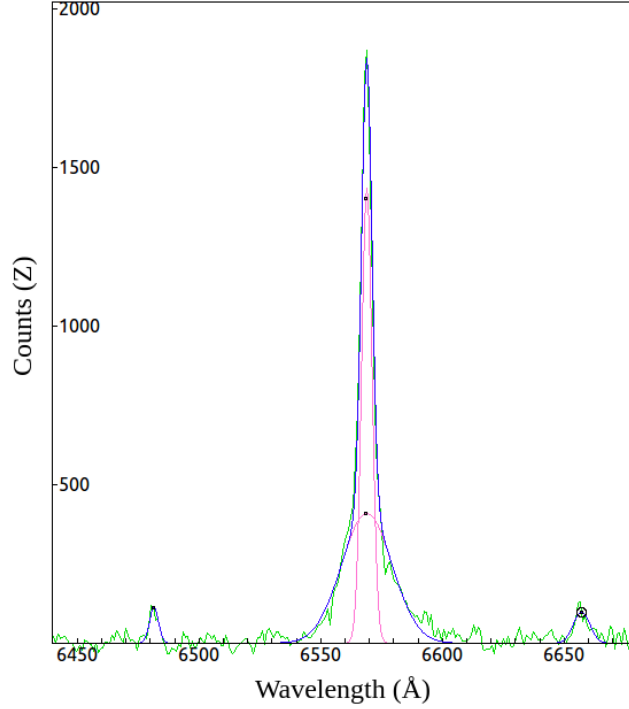


Fig. 5.18: Model fit to the RXJ0513 H α emission observed on 27 August 2012. A combination of four Gaussian profiles was used: one for the main emission line, one for the broad emission wing, and one for each of the satellite features associated with the jet. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

Table 5.15: The ratio of the equivalent widths of the main components of the He II λ 4686 and H α emission lines in RXJ0513 obtained from the SALT observations on different dates.

Observation date	He II λ 4686 main / H α main
27 Aug 2012	0.294 ± 0.071
29 Oct 2012	0.259 ± 0.068
18 Sep 2013	0.44 ± 0.11
22 Sep 2013	0.442 ± 0.090
25 Sep 2013	0.60 ± 0.22

features were each fitted with a Gaussian profile as well, as shown in Fig. 5.19. A single Gaussian profile was fitted to the main He II λ 5411 emission line component in all the spectra, with an additional Gaussian profile for its wing structure for the observation on 27 Aug 2012.

By utilizing the absolute values of the equivalent widths obtained from the Gaussian fits, line ratios were calculated for RXJ0513. The only line ratio in which variability outside the 99.73% error margins was observed, is the ratio of the main He II λ 4686 to the main H α emission. This ratio was higher in Sep 2013 than in 2012, as shown in Table 5.15. The mean of each line ratio over all the RXJ0513 observations was also calculated, and the results are presented in Table 5.16.

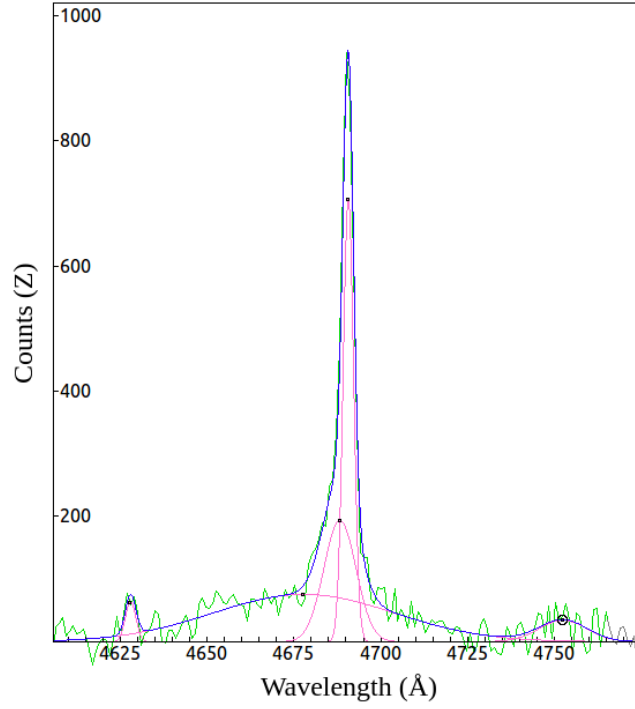


Fig. 5.19: Model fit to the RXJ0513 He II $\lambda 4686$ emission observed on 27 August 2012. A combination of five Gaussian profiles was used: one for the main emission line, one for the broad emission wing, one for the C III, C IV, N III, Si IV blend, and one for each of the satellite features associated with the jet. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

Table 5.16: Mean line ratios of the features in the spectra of RXJ0513, calculated from the absolute values of their respective equivalent widths, and averaged over all the SALT observations. The equivalent width of the He II $\lambda 4686$ wing includes the contribution of the C III, C IV, N III, Si IV blend.

Features compared	Mean line ratio
H α wing / H α main	1.29 ± 0.28
H α blue satellite / H α main	0.063 ± 0.062
H α red satellite / H α main	0.11 ± 0.12
H β main / H α main	0.251 ± 0.043
H γ main / H α main	0.164 ± 0.057
He II $\lambda 4686$ main / H α main	0.407 ± 0.055
He II $\lambda 5411$ main / H α main	0.128 ± 0.05
H β wing / H β main	1.3 ± 1.2
H β blue satellite / H β main	0.12 ± 0.29
H β red satellite / H β main	0.23 ± 0.7
H γ main / H β main	0.69 ± 0.3
He II $\lambda 4686$ main / H β main	1.69 ± 0.4
He II $\lambda 5411$ main / H β main	0.53 ± 0.24
H γ wing / H γ main	2.2 ± 3.7
He II $\lambda 4686$ main / H γ main	2.69 ± 0.85
He II $\lambda 5411$ main / H γ main	0.9 ± 0.41
He II $\lambda 4686$ wing / He II $\lambda 4686$ main	1.63 ± 0.86
He II $\lambda 4686$ blue satellite / He II $\lambda 4686$ main	0.09 ± 0.15
He II $\lambda 4686$ red satellite / He II $\lambda 4686$ main	0.29 ± 0.86
He II $\lambda 5411$ main / He II $\lambda 4686$ main	0.33 ± 0.12
He II $\lambda 5411$ wing / He II $\lambda 5411$ main	2.1 ± 9.0

To compare the evolution of the spectral properties of RXJ0513 with the variability in the overall optical flux, the W_λ and FWHM values from Tables 5.11, 5.12, 5.13 and 5.14 were plotted together with a portion of the OGLE-IV lightcurve in Fig. 5.20 and Fig. 5.21 respectively. The OGLE lightcurves will be discussed in more detail in §6.2.

Since the highly blue- and redshifted satellite features in $H\alpha$, $H\beta$ and $\text{He II } \lambda 4686$, as well as the broad wing of $\text{He II } \lambda 5411$, were only observed and fitted for a single observation on 27 Aug 2012, they are not included in Fig. 5.20 and 5.21. Also, since the $\text{He II } \lambda 4686$ wing is blended with the C III , C IV , N III and Si IV emission, its FWHM is not included in Fig. 5.21, as it is not really physically meaningful.

The 27 Aug and 29 Oct 2012 observations were obtained directly before and directly after an optical low state, while the Sep 2013 observations were during an optical high state. Unfortunately, no SALT observations of RXJ0513 were obtained during an optical low state. It is evident from Fig. 5.20 that no statistically significant variability was revealed in the equivalent widths of the wing structures of any of the lines. However, a decrease was observed in W_λ for the main $H\alpha$ and $H\beta$ emission components during the on-state from 22 to 25 Sep 2013. There were no simultaneous OGLE pointings, but judging from the OGLE data points just before 22 Sep and after 25 Sep, there might have been a simultaneous, very slight decrease in the overall optical brightness. On 22 Sep 2013, W_λ for the $\text{He II } \lambda 4686$ line was higher than in the 2012 observations bordering the optical low state.

Considering the FWHM values in Fig. 5.21, the values for the $\text{He II } \lambda 4686$ and $H\alpha$ main components, as well as the wing of the latter, were larger in Sep 2013 than during the first observation (27 Aug 2012). However, the $H\alpha$ wing FWHM exhibited a peculiar dip from 18 to 22 Sep, increasing again on 25 Sep. The main $H\beta$ component also had a larger FWHM on 18 Sep, on a marginally significant level. In general, the FWHM values of the main components seemed to decrease slightly from 18 to 25 Sep, with a presumed simultaneous decrease in the OGLE optical flux.

5.7.2 Discussion

The SALT optical spectra of RXJ0513 exhibit similar characteristics to those that have been obtained for the source since its discovery more than two decades ago. The spectra are also very similar to the CAL 83 spectra, with the dominant emission features being those of H I and He II , with weaker emission of ionized carbon, nitrogen and oxygen also present. The mean line ratios for RXJ0513 as listed in Table 5.16 are similar to those measured for the same lines in CAL 83 (Table 5.9), with no differences outside the 99.73% confidence intervals.

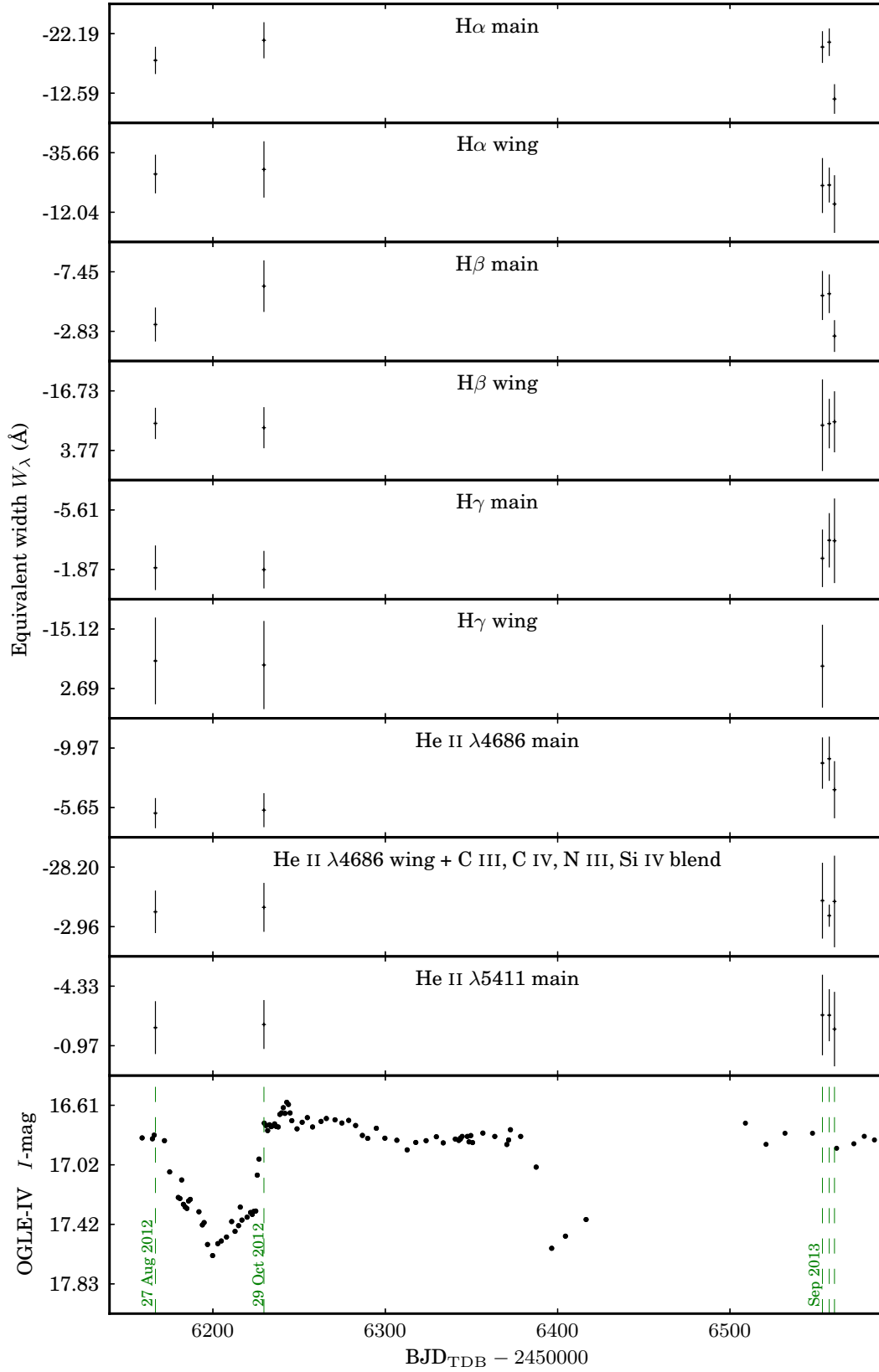


Fig. 5.20: The equivalent widths of spectral features of RXJ0513 measured with SALT, as listed in Tables 5.11, 5.12, 5.13 and 5.14, plotted together with the OGLE-IV lightcurve. The W_λ errors represent 99.73% confidence levels obtained by fitting Gaussian profiles. The OGLE-IV errors were also plotted, although they are mostly smaller than the data points.

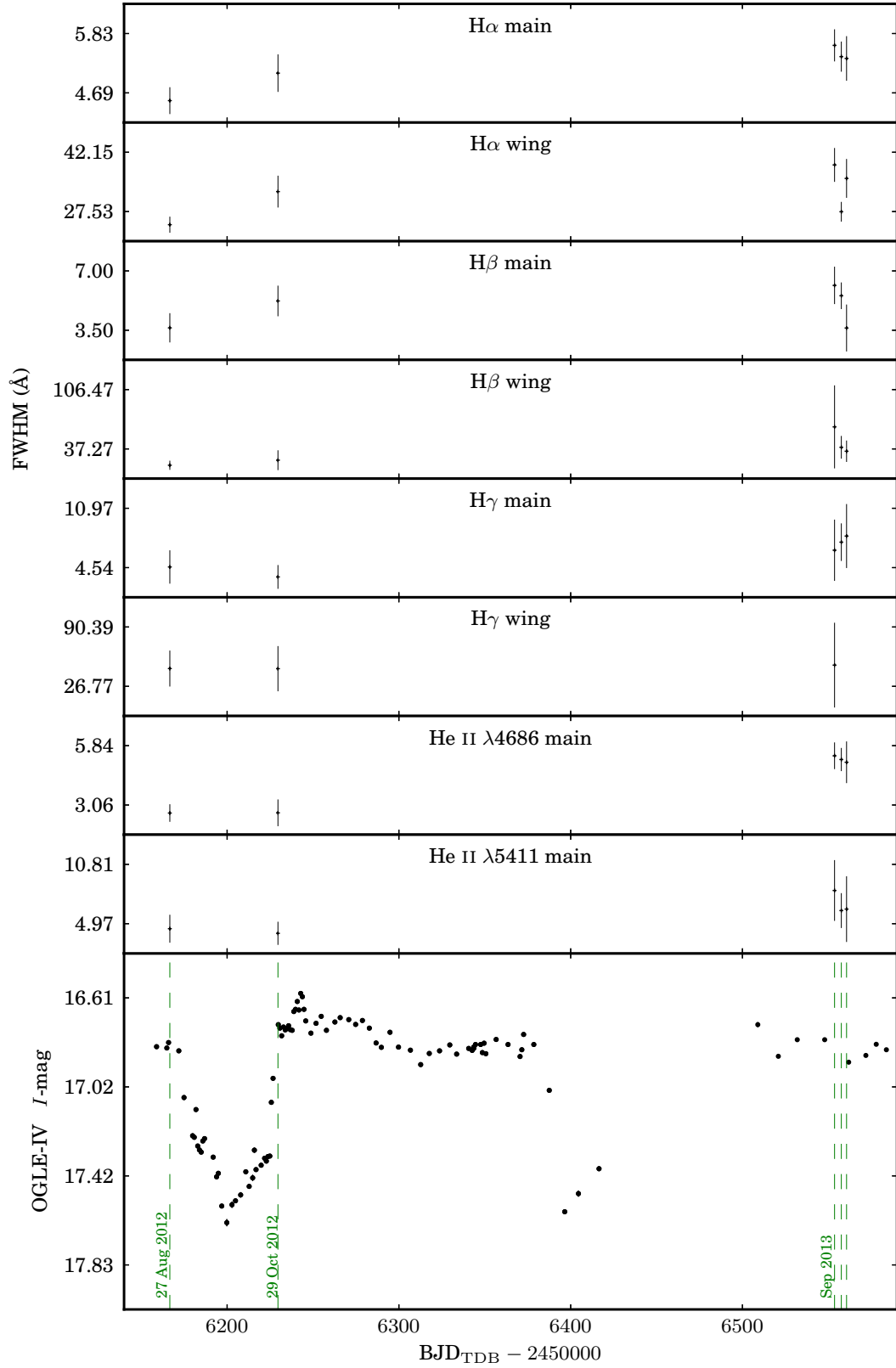


Fig. 5.21: The FWHM values of the RXJ0513 spectral features measured with SALT, as listed in Tables 5.11, 5.12, 5.13 and 5.14, plotted together with the OGLE-IV lightcurve. The FWHM errors represent 99.73% confidence levels obtained by fitting Gaussian profiles. The OGLE-IV errors were also plotted, although they are mostly smaller than the data points.

The main difference between the spectra of RXJ0513 and those of CAL 83 is the occurrence of sharp bipolar satellite features of Balmer and He II in RXJ0513. These S^- and S^+ features were fitted for the $H\alpha$, $H\beta$ and He II $\lambda 4686$ lines in the 27 Aug 2012 spectra, and their positions relative to the respective central emission components are approximately -4000 km s^{-1} and $+4000 \text{ km s}^{-1}$ (see Tables 5.11, 5.12 and 5.14). Although the error bars on the weak satellite feature measurements are quite large, it appears as if both the W_λ and FWHM of S^+ are considerably larger than that of S^- for all 3 of the lines. The line with the best S/N is $H\alpha$, so the current discussion will focus on the satellite features associated with this line. As shown in the previous section, the satellite features are variable, and the S^- feature is sometimes detected in absorption.

The satellite features most probably originate from biconical plasma jets, with the S^- component being formed in the jet component approaching the observer, launched from the side of the accretion disc facing the observer. The other jet component may be launched from the opposite side of the accretion disc, and since it is receding from the observer, it produces the S^+ component. This is similar to the twin jet configuration in the well known Galactic microquasar (radio X-ray binary) SS 433, although the velocities in this source are about an order of magnitude larger than in the SSSs (e.g. Blundell, Bowler and Schmidtobreick, 2007). Although the region in the receding jet close to the opposite side of the disc will be obscured from the observer, the observer would be able to see the optically thin emission from further out in the receding jet extending from behind the accretion disc, provided that the angle between the jet axis and the line of sight is non-zero.

If the blueshifted S^- feature from the approaching jet is observed in absorption, it implies that emission from the central continuum source is absorbed by optically thin material in the approaching jet, i.e. the continuum source is behind the jet along the line of sight. When the S^- component is in emission, it implies that the optically thin emission from the jet is observed along a line of sight that does not intersect the continuum source. During the orbital motion of the WD, it is possible that the S^- component may be observed in absorption at some orbital phases where it happens to obscure the continuum source, and in emission at other phases where it does not. This would depend on the orbital inclination, the orientation of the twin jet axis, and the actual location of the dominant source of optical emission in the system, which is most likely the accretion disc rim. An investigation of this scenario should form part of follow-up work.

It has been shown that the mass of the WD in RXJ0513 is probably $\sim 1.3 M_\odot$, similar to the mass of the WD in CAL 83 (§3.2). The corresponding WD radius according to Eq. (2.25) is $\sim 5 \times 10^8 \text{ cm}$. According to the two-temperature model

fitted by Burwitz et al. (2007) to the *Chandra* LETGS spectrum of RXJ0513, the bolometric luminosity of the source is $\sim 8 \times 10^{37}$ erg s $^{-1}$ (see Table 3.6). The upper limit on the accretion rate below which irradiation-induced disc warping is possible, is given by Eq. 2.49 in §2.4.4 as

$$\dot{m}_{\text{acc}} \lesssim 6 \times 10^{-8} \left(\frac{R_{\text{disc}}}{100 R_{\text{wd}}} \right)^{1/2} \left(\frac{L_{\text{nuc}}}{8 \times 10^{37} \text{ erg s}^{-1}} \right) \left(\frac{R_{\text{wd}}}{5 \times 10^8 \text{ cm}} \right)^{1/2} \left(\frac{M_{\text{wd}}}{1.3 M_{\odot}} \right)^{-1/2} M_{\odot} \text{ yr}^{-1}. \quad (5.4)$$

If the accretion rate in RXJ0513 is as high as $\sim 10^{-6} M_{\odot} \text{ yr}^{-1}$, as has been proposed previously (see §3.2.3), disc warping is therefore highly unlikely to be the cause of the changing structures in the bipolar jet features in the optical spectra.

Although the orbital inclination of RXJ0513 is not known, it has been estimated by several authors to be quite low. It has been shown in §5.6.2 that the escape velocity associated with a massive WD is of the order of 6000 km s $^{-1}$. If the inclination angle in RXJ0513 is low, the observed bipolar velocities of ± 4000 km s $^{-1}$ constitute a substantial fraction of v_{esc} .

In addition to the bipolar features, the emission lines also exhibit broad bases somewhat similar to the broad wing structures observed in CAL 83. The main Balmer emission line structure is asymmetric, probably due to blue absorption. Two slightly different types of outflows may thus be present in RXJ0513: a well-collimated bipolar jet, as well as a wind with lower associated velocities.

The widths of the central main components of the emission lines in RXJ0513 are similar to those of CAL 83. It has been shown in §5.6.2 that the velocity dispersion of these components may arise because of rotational line broadening in an accretion disc. The equivalent widths and FWHM values of the emission lines as shown in Fig. 5.20 and Fig. 5.21 also exhibit small variations which seem to be approximately correlated with the overall optical flux, as was seen in CAL 83.

The possibility of an anti-correlation between the intensity of the O VI $\lambda 5290$ line and the optical brightness, as seen in CAL 83, has also been investigated in RXJ0513. This feature, as well as the O V $\lambda 5598$ and O VI $\lambda \lambda 3811, 3834$ lines was faintly visible, but very weak, in all the SALT observations. Using the intensity of the nearby He II $\lambda 5411$ line as a qualitative reference, shows the O VI $\lambda 5290$ emission to be significantly fainter than this He II line in all the SALT observations. However, no SALT observations were obtained during an optical low, X-ray high state.

Consequently, the behaviour of the O VI $\lambda 5290$ line in previously published spectra as summarized in §3.2.3 was evaluated. The average spectrum in Fig. 3.38 was obtained by Crampton et al. (1996) at the CTIO 4-m telescope during 5 consecutive nights starting on 10 November 1994 ($\text{BJD}_{\text{TDB}} \gtrsim 2449666$). According to the MACHO lightcurve presented in Fig. 6.7 in §6.2, RXJ0513 was in an optical high state at this time. The intensity of O VI $\lambda 5290$ with respect to He II $\lambda 5411$ is qualitatively comparable to the relation observed with SALT during optical high states.

An even smaller ratio of O VI $\lambda 5290$ to He II $\lambda 5411$ was observed by Cowley et al. (1998) with the CTIO 4-m telescope in November 1996 ($\text{BJD}_{\text{TDB}} \sim 2450392$) (see Fig. 3.14), again during an optical high state according to the MACHO lightcurve. Observations of the spectral regions containing the O VI $\lambda 5290$ line are therefore needed during an optical low state to evaluate the behaviour of the O VI line when the WD envelope is in its contracted state.

5.8 MR Vel

5.8.1 Results

The calibrated, continuum-subtracted spectra of MR Vel obtained during the observations listed in Table 5.3 are shown in Fig. 5.22, with the detected emission lines marked on each spectrum. The properties of all the emission lines that have been detected previously in MR Vel are listed in Table 5.17 (see §3.3.2). For reference, their positions are also indicated in the spectra in Fig. 5.22, regardless of whether or not they were observed in the respective spectra. The line labels in a particular spectrum were offset by a fixed velocity equal to the mean velocity of the fitted lines in that particular spectrum.

The main features in the spectra are the He II $\lambda 4686$ and H α (possibly blended with He II $\lambda 6560$) lines, while He II $\lambda 4541$, H β and He II $\lambda 5411$ are also detected. An emission feature redshifted by $\sim 5400 \text{ km s}^{-1}$ relative to H α may be an H α satellite feature associated with a jet. An equivalent blueshifted component was not detected in the SALT spectra. The He II $\lambda 4686$ and H α lines observed on 21 Mar 2013 are plotted in terms of velocity relative to the central line wavelength in Fig. 5.23.

As in the case of CAL 83 and RXJ0513, ionized carbon, nitrogen and oxygen emission lines are also present. The C III, C IV, N III blend at $\sim 4660 \text{ \AA}$ is evident in the SALT spectra, while O VI $\lambda 5290$ and C IV $\lambda \lambda 5801, 5812$ may be faintly visible above the noise level. The [Fe XIV] $\lambda 5303$ line proposed by Schmidtke et al. (2000) might be present.

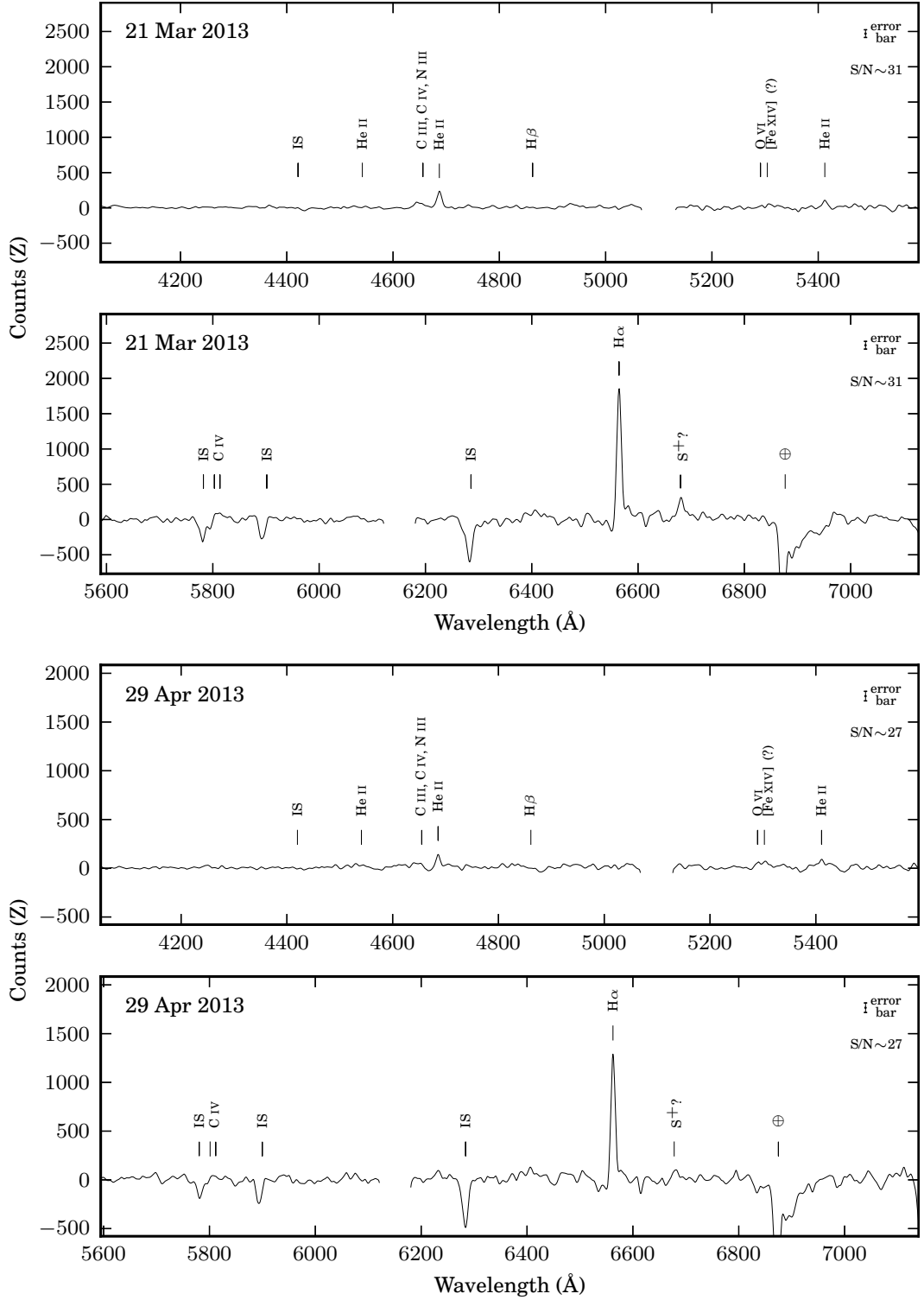


Fig. 5.22: SALT RSS spectra of MR Vel. The vertical axis represents the counts after subtraction of the fitted continuum. The gaps in the spectra are due to the gaps between the CCD chips. The $\oplus$ symbol is used to indicate the atmospheric A- and B-bands. “IS” refers to the interstellar absorption bands. The vertical error bar (upper right corner) indicates the typical 1σ noise amplitude about a data point, and the estimated S/N of the spectrum is also given. These quantities were determined with the method in §5.5.3.

Table 5.17: Standard properties of the emission lines observed in MR Vel. The third column indicates the temperature range within which the ion can exist in the relevant ionization state without being excited to an even higher state, according to the ionization potentials and assuming thermal ionization. The last column represents the thermal broadening FWHM according to Eq. (C.11) for the single lines, corresponding to the highest temperature in the third column. The rest wavelengths and ionization potentials were obtained from the NIST Atomic Spectra Database.

Element and ionization state	Rest wavelength λ_0 (Å)	Temperature range (K)	Thermal broadening FWHM (Å)
H I (H α) with bipolar satellite features (Blended with He II λ 6560)	6562.79	<157700	1.9
H I (H β) (Blended with He II λ 4859)	4861.35	<157700	1.4
He II	4541	285200-631200	1.3
He II	4686	285200-631200	1.3
He II	5411	285200-631200	1.5
Blend of C III, C IV, and N III emission	$\sim$ 4630-4680	282900-555500 555500-748100 343400-550400	
C IV	5801.33	555500-748100	1.0
C IV	5811.98	555500-748100	1.0
O VI	5290	1321000-1602000	1.2
Fe XIV	5302.86	4188000-4550000	1.1

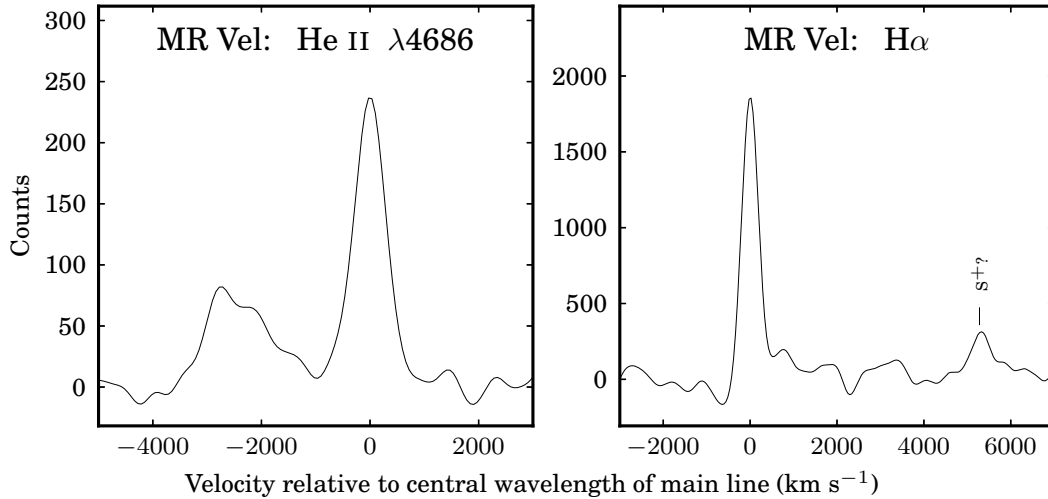


Fig. 5.23: The spectral regions around the He II λ 4686 and H α emission lines in the MR Vel spectrum on 21 Mar 2013, plotted in terms of velocity relative to the central line position. Note the Bowen complex on the blue side of He II λ 4686. The H α line exhibits a P Cyg profile, and also a possible redshifted satellite feature at approximately +5400 km s $^{-1}$.

Table 5.18: MR Vel emission lines: spectral line fitting results. The quoted velocity shift v_{SSB} is with respect to the rest wavelength for the main components, but for the absorption and satellite components it is with respect to the central wavelength of the relevant main component.

Date	BJD _{TDB} −2450000 (mid-exposure)	v_{SSB} (km s ^{−1})	FWHM (Å)	FWZI (Å)	W_{λ} (Å)
Hα main component					
21 Mar 2013	6373.46991	64 $\pm$ 11	5.99 $\pm$ 0.58	20.5	−4.0 $\pm$ 1.1
29 Apr 2013	6412.36561	−32 $\pm$ 10	5.74 $\pm$ 0.72	19.5	−3.4 $\pm$ 1.1
Hα blueshifted absorption					
21 Mar 2013	6373.46991	−590 $\pm$ 160	9.4 $\pm$ 8.9	30.3	0.5 $\pm$ 1.4
29 Apr 2013	6412.36561	−730 $\pm$ 580	27 $\pm$ 27	85.0	0.9 $\pm$ 4.4
Hα redshifted jet?					
21 Mar 2013	6373.46991	5316 $\pm$ 82	9.6 $\pm$ 4.2	31.3	−0.8 $\pm$ 1.5
29 Apr 2013	6412.36561	5390 $\pm$ 190	10.0 $\pm$ 9.9	33.2	−0.4 $\pm$ 1.7
He II λ4686					
21 Mar 2013	6373.46991	87 $\pm$ 42	7.0 $\pm$ 1.6	22.5	−7.5 $\pm$ 1.3
29 Apr 2013	6412.36561	−3 $\pm$ 61	4.5 $\pm$ 2.2	14.6	−4.33 $\pm$ 0.92
C III, C IV, N III blend					
21 Mar 2013	6373.46991		19.9 $\pm$ 9.7	64.5	−5.7 $\pm$ 3.1
29 Apr 2013	6412.36561		20 $\pm$ 15	63.5	−5.7 $\pm$ 3.5

Detailed fitting as described in §5.5.4 and §5.5.5 was only performed for the strongest lines in each spectrum, and the results are presented in Table 5.18. The H α emission was fitted with 3 Gaussian components: one for the main emission, one for the blueshifted absorption, and one for the possibly redshifted satellite feature (S⁺) (e.g. Fig 5.24). The He II λ 4686 emission line and the C III, C IV, N III blend were fitted simultaneously with a combination of two Gaussian profiles.

It is evident from Table 5.18 that no significant variability was detected in the FWHM values of any of the features between the observations of 21 Mar and 29 Apr 2013. The equivalent widths also did not vary significantly, except for the He II λ 4686 line, where W_{λ} varied from $−7.5 \pm 1.3$ Å to $−4.33 \pm 0.92$ Å.

From the equivalent width values, the line ratios for the different features in the spectrum were calculated. No significant variability (i.e. outside of the 99.73% margins) was detected in the line ratios between the observations of 21 Mar and 29 Apr 2013. Therefore, only the mean values of these ratios are provided in Table 5.19.

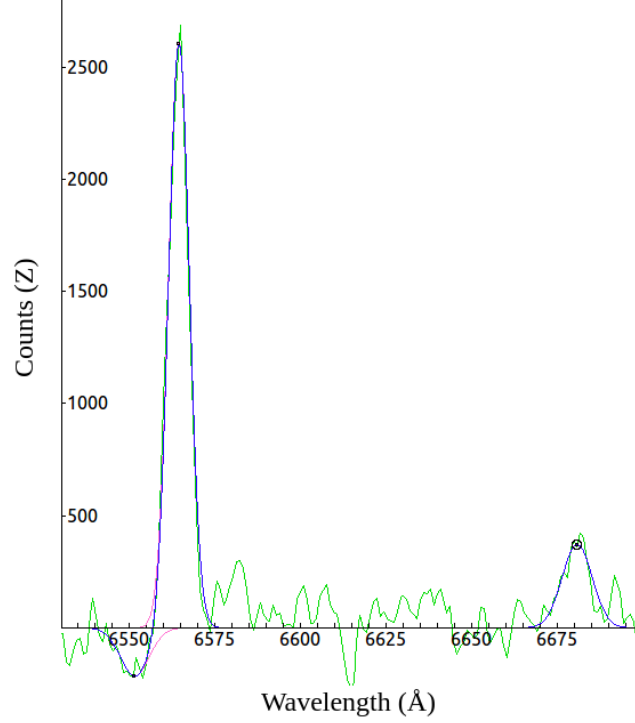


Fig. 5.24: Model fit to the MR Vel $H\alpha$ spectral line observed on 21 Mar 2013. Three Gaussian profiles were used: one for the main emission line, one for the blueshifted absorption, and one for the redshifted, probably jet-related feature. Green represents the continuum-subtracted data, pink lines indicate the individual Gaussian profiles, and blue lines the total model.

Table 5.19: Mean line ratios of the features in the spectra of MR Vel, calculated from the absolute values of their respective equivalent widths, and averaged over both the SALT observations.

Features compared	Mean line ratio
$H\alpha$ absorption / $H\alpha$ main	0.20 ± 0.66
$H\alpha$ red satellite / $H\alpha$ main	0.15 ± 0.31
He II $\lambda 4686$ / $H\alpha$ main	1.58 ± 0.39
C III, C IV, N III blend / He II $\lambda 4686$ main	1.04 ± 0.48

5.8.2 Discussion

The spectra of MR Vel share many characteristics with those of CAL 83 and RXJ0513, with emission lines of mainly H I and He II, but also weaker C, N and O emission. It was shown in the previous section that no significant change in the spectrum was observed from one SALT observation to the next, apart from the decrease in W_λ of He II $\lambda 4686$. The spectral properties can now be compared to the previously published observations in §3.3.2.

Motch et al. (1994) commented on the existence of narrow and broad components in the H α and He II $\lambda 4686$ lines, and also the occasional presence of double-peaked structures in these lines. These structures were not observed in the SALT spectra, although a faint red wing may be present in H α .

Blue and red bipolar satellite features in H α were observed on 7 June 1998 by Motch (1998), shifted by ± 5200 km s $^{-1}$ relative to the main H α emission. The blue jet component has not been observed since then. On 8 June, a feature still present at ~ 6680 Å was ascribed to He I $\lambda 6678$ emission by Motch (1998). This feature was observed again in the spectra of Matsumoto and Mennickent (2000) and Schmidtke et al. (2000). Schmidtke et al. considered it to be unlikely that the feature at ~ 6680 Å is He I $\lambda 6678$ or He II $\lambda 6683$ emission, as no other He I lines are present in the spectrum, and the position, intensity and width of the feature are not consistent with the other He II lines in the spectrum.

Schmidtke et al. (2000) therefore suggested that the ~ 6680 Å feature represents the redshifted S $^+$ satellite of H α with a velocity of ~ 5350 km s $^{-1}$, which may be visible independently of the simultaneous presence of the corresponding blueshifted S $^-$ feature. This is then similar to what has been observed in RXJ0513, where the S $^-$ feature is sometimes absent or in absorption. Schmidtke et al. proposed that the visibility of S $^-$ emission may be dependent on orbital phase.

The ~ 6680 Å feature was also observed in the SALT spectra, and is thus likely to represent emission from a collimated jet with mean velocity 5350 ± 100 km s $^{-1}$ (see Table 5.18). This is of the same order as the escape velocity of a WD. The P Cyg profile also observed in H α serves as additional evidence of outflows. The fit values of the absorption component were better constrained on 21 Mar 2013 than on 29 Apr 2013, with its centre displaced by approximately -590 km s $^{-1}$ with respect to the main emission line. The FWZI of the H α absorption is ~ 1400 km s $^{-1}$, i.e. the absorption blends into the continuum at $(-590 - \frac{1}{2} \times 1400)$ km s $^{-1} \sim -1300$ km s $^{-1}$. The terminal velocity of the outflow in which the absorption component originates is therefore approximately 1300 km s $^{-1}$, i.e. a lower velocity outflow component may

be present in addition to the well-collimated jet.

Motch et al. (1994) observed variability in the equivalent width of He II $\lambda 4686$ within the -4 to -9 Å range. It was later shown by Schmidtke et al. (2000) that the equivalent widths (and also the radial velocities) of H α and He II $\lambda 4686$ vary in phase with the V -magnitude during the course of the orbital motion. The variation was ~ 1 Å for He II $\lambda 4686$, and slightly smaller for H α . The change in the He II $\lambda 4686$ W_λ from -7.5 ± 1.3 Å to -4.33 ± 0.92 Å according to the SALT data is in agreement with these findings.

Apart from the above small variations in the emission lines, a comparison of the line widths and intensities published since 1994 reveals no additional long-term changes in the line emission from the system. To date, no observational evidence exists that MR Vel undergoes any X-ray off-states or drastic changes in its optical brightness like CAL 83 and RXJ0513. It may therefore truly be a “steady supersoft source” according to the canonical model, or its cycles may simply be much longer than the existing observational baseline.

In this chapter, the properties of the spectral lines of CAL 83, RXJ0513 and MR Vel in the optical waveband have been investigated. In the next chapter, photometric measurements will be utilized to investigate the variability of the broadband optical emission from these sources on various time-scales.

Chapter 6

Optical photometry

The detection of dwarf nova oscillations and quasi-coherent oscillations in the optical lightcurves of cataclysmic variables, as well as the presence of a ~ 67 s, possible DNO-type oscillation in the X-ray lightcurves of CAL 83, raises the question as to whether similar modulations can be found in the optical lightcurves of SSSs. The *XMM-Newton* Optical Monitor data were obtained in imaging mode, and thus did not provide the required timing resolution. To address this topic, fast photometric observations of CAL 83, RXJ0513 and MR Vel have been performed with the sensitive Sutherland High-Speed Optical Cameras (SHOC) on the SAAO 1.9-m telescope, using exposure times up to only a few seconds. These results are reported in §6.1 of this chapter.

An important aspect to keep in mind is that the dominant source of optical emission in these SSSs is the irradiated accretion disc. In fact, the secondary star is not detected at all in the optical spectra. Any modulations in the optical flux is therefore directly linked to variability in the accretion disc, or, to a lesser extent, in the reprocessed radiation from the heated side of the secondary. X-ray emission from regions close to the WD may also be observed at a smaller amplitude in the optical waveband, either simply due to the tail of the thermal emission extending to lower energies, or due to the scattering and/or reprocessing of radiation in the disc.

The superorbital modulations in the optical lightcurves of CAL 83 and RXJ0513 on time-scales of hundreds of days, and the anti-correlated X-ray flux, are also key elements in understanding the accretion modes occurring in these sources. The optical monitoring of these sources as part of the MACHO (MASSive Compact Halo Objects) and OGLE (Optical Gravitational Lensing Experiment) Projects, allowed the study of these cycles in unprecedented detail. In §6.2, an update is provided on the work of previous authors by presenting the most recent OGLE-IV lightcurves.

6.1 Photometry with SHOC on the SAAO 1.9-m

6.1.1 The SAAO 1.9-m telescope

This telescope is located at the South African Astronomical Observatory (SAAO), which is approximately 18 km east of Sutherland in the Northern Cape, with coordinates $20^{\circ} 48' 38.5''$ E; $32^{\circ} 22' 46''$ S, and an altitude of ~ 1800 m. It was located at the Radcliffe Observatory in Pretoria until 1974, and is therefore sometimes referred to as the Radcliffe telescope. The parabolic primary mirror has a diameter of 1.9 m, (74 inches), and offers both Cassegrain and Newtonian configurations for the secondary focus. The focal ratio of the beam at the Cassegrain focus is $f/18$, with a plate scale of ~ 6 arcsec/mm. More information related to this telescope can be found on the SAAO website¹.

6.1.2 The Sutherland High-Speed Optical Cameras (SHOC)

The Sutherland High-Speed Optical Cameras (SHOC) have been commissioned for use on the 1.9-m, 1.0-m and 0.75-m telescopes at the Sutherland site during 2013. The two cameras (SHOC 1 and SHOC 2) are nearly identical, with only small differences in their technical properties, e.g. read noise and saturation limits. Either one of these cameras can be mounted at the Cassegrain focus of the 1.9-m telescope. On this telescope, SHOC can be operated either with or without a focal reducer, yielding a field of view of $2.79 \text{ arcmin} \times 2.79 \text{ arcmin}$ or $1.29 \text{ arcmin} \times 1.29 \text{ arcmin}$ respectively. A detailed discussion of these cameras can be found in the instrument paper (Coppejans et al., 2013), on the SHOC web page² and in the user manual, SHOCnhelpful³ (Coppejans and Gulbis, 2014), therefore only a few key properties will be mentioned here.

Each SHOC system consists of a back-illuminated Andor iXon X3 888 UVB camera, a global positioning system (GPS), a control computer and peripherals. The GPS, computer and the power supply of the camera are built into a single “control crate” unit. Three such control units exist: shocnawe, shocndisbelief and shocnhorror. Therefore SHOC 1 and SHOC 2 can be operated simultaneously at different telescopes, with one control unit serving as a backup.

Each CCD has 2048×1024 pixels and is operated in frame transfer mode, i.e. with an imaging area of 1024×1024 . The CCD pixel size is $13 \mu\text{m}$, yielding a plate scale of 0.163 arcsec/pixel with the focal reducer, and 0.076 arcsec/pixel without. The CCDs are thermoelectrically cooled to a typical operating temperature of -60°C ,

¹<http://www.sao.ac.za/science/facilities/telescopes/1-9m/>

²<http://www.sao.ac.za/science/facilities/instruments/shoc/>

³<http://shoc.sao.ac.za/Documents/ShocnHelpful.pdf>

with a dark current of $<0.001 \text{ e}^- \text{ pixel}^{-1} \text{ s}^{-1}$. The cameras allow several readout modes: conventional (CON) mode with a horizontal pixel shift speed of 1 or 3 MHz, as well as an electron multiplying (EM) mode with shift speeds of 1, 3, 5 or 10 MHz. The 1 MHz amplifiers make use of a 16-bit analog-to-digital converter (ADC), while the faster modes use a 14-bit ADC, and therefore have a smaller dynamic range.

SHOC is optimized for high-speed photometry at visible wavelengths. Its quantum efficiency is $<35\%$ at wavelengths $<3750 \text{ \AA}$, $>90\%$ between 4800 and 7000 $\text{\AA}$ and $<40\%$ at wavelengths longer than 9000 $\text{\AA}$. During normal operation, it can achieve frame rates between 1 and 20 frames s^{-1} , and up to ~ 100 frames s^{-1} with binning and subframing. Individual frames are triggered by one of three methods: (i) internal, (ii) external or (iii) external start.

In internal mode, the internal computer clock is used to trigger the frames and to write time stamps to the file headers. The computer clock drifts by approximately 1 s per day relative to the GPS. Therefore, if the computer is synced to the GPS at the beginning of the night, the absolute timing in internal mode is only ~ 1 s. In external mode, each frame is triggered using the GPS, while in external start mode, only the first frame is triggered with the GPS, with subsequent frames triggered internally. According to the Andor specifications, the absolute timing accuracy of GPS triggered frames should be $\sim 10 \text{ }\mu\text{s}$.

By utilizing a Light Pulse Generator (LPG), Kotze (2014) undertook a detailed investigation into the timing accuracy of the different triggering modes, and her findings were the following: (i) In addition to the ~ 1 s uncertainty in the absolute timing in internal mode, it was found that there is also a $\sim 0.00001\text{--}0.00002$ s uncertainty in the relative timing, due to the uncertainty in the 5th decimal of the exposure time and dead time. This uncertainty can accumulate up to ~ 2 s after 100000 frames. (ii) The investigation of Kotze (2014) confirmed that the absolute and relative timing for GPS triggered frames in external mode is accurate to within the fastest sampling rate of 10 ms per exposure. (iii) The absolute timing accuracy of the first frame in external start mode is very good, but because internal triggering is used for the subsequent frames, the relative timing also has an uncertainty in the 5th decimal. However, the relative timing error in external start mode is 0.000019 s, and is not identical to that observed in internal mode. Therefore, for high time resolution studies requiring $\lesssim 1$ s accuracy, external mode is recommended.

6.1.3 Observations with SHOC

SHOC observations of CAL 83, RXJ0513 and MR Vel were obtained in April 2013 and December 2014. The details pertaining to these observations are given in Ta-

Table 6.1: SHOC observation sequences of CAL 83, RXJ0513 and MR Vel. See Table 6.2 for a description of the instrumental configurations.

Target	Observation number	Start date & time (UT)		Exposures				Configuration
CAL 83	01	2013-04-10	19:17	1 s	×	549 =	549 s	A
	02	2013-04-11	20:01	5 s	×	718 =	3590 s	B
	03	2013-04-12	17:50	2 s	×	2998 =	5996 s	B
	04		19:32	2 s	×	2260 =	4520 s	B
	05	2013-04-13	17:42	1.5 s	×	2598 =	3897 s	B
	06	2013-04-14	18:28	0.5 s	×	3998 =	1999 s	B
	07	2013-04-15	17:53	2 s	×	1618 =	3236 s	B
	08	2014-12-16	20:43	5 s	×	298 =	1490 s	C
	09		21:14	10 s	×	718 =	7180 s	C
	10		23:20	8 s	×	718 =	5744 s	C
	11	2014-12-17	00:59	8 s	×	673 =	5384 s	C
RXJ0513	01	2013-04-13	19:05	3 s	×	1198 =	3594 s	B
	02	2013-04-14	19:15	1.5 s	×	1798 =	2697 s	B
	03	2014-12-30	18:55	10 s	×	1438 =	14380 s	C
MR Vel	01	2013-04-11	21:15	1.5 s	×	546 =	819 s	B
	02		21:32	0.1 s	×	65534 =	6553 s	B
	03		23:28	0.1 s	×	6463 =	646 s	B
	04	2013-04-13	21:14	0.2 s	×	23398 =	4680 s	B
	05		22:36	0.2 s	×	10498 =	2100 s	B
	06	2013-04-14	21:12	0.06 s	×	59998 =	3600 s	B
	07		22:20	0.0558 s	×	59998 =	3348 s	B
	08	2013-04-15	20:49	0.4 s	×	5019 =	2008 s	B

ble 6.1, and the relevant instrumental configurations in Table 6.2. Because the primary goal was to search for optical variability on time-scales of ~ 1 s up to several minutes, a clear filter was used for all the observations to maximize the signal-to-noise ratio (S/N).

A set of 30 evening twilight sky flats was obtained with a clear filter and the same settings as the target frames almost every day, with the following exceptions: On 10 Apr 2013, only 20 flat frames were obtained. No flat fields were obtained on 14 and 15 Apr 2013, so the flats of 13 Apr 2013 were used to calibrate the target frames obtained on the 14th and 15th. A set of 30 bias frames with the same binning as the target frames were also obtained every day, except on 14 Apr 2013, for which the bias frames recorded on 13 Apr 2013 were used.

6.1.4 Data reductions

The SHOC data is stored in the form of data cubes, i.e. each series of exposures of a single target is stored in a single FITS file (see §5.3.1). This file contains a single HDU with a header applying to the whole cube, and all the consecutive im-

Table 6.2: Instrumental configurations used for the SHOC observations of CAL 83, RXJ0513 and MR Vel listed in Table 6.1. CON mode was utilized for all observations, using a clear filter in both filter wheels.

Configuration	Camera	Pixel shift speed (mHz)	Gain	Binning	Observer
A	SHOC2	3	$5.2\times$	16×16	Alida Odendaal
B	SHOC2	3	$5.2\times$	8×8	Alida Odendaal
C	SHOC1	1	$2.5\times$	8×8	Lizelke Klindt

ages stored in the data unit. The data reductions were performed by utilizing the Python- and PyRAF-based SHOC data reduction pipeline developed by Marissa Kotze (the version updated on 3 June 2014). The pipeline, as well as a detailed description of its operation, is available online⁴.

As input to the pipeline, the user provides the 3 cubes containing the target, flat and bias frames respectively. The target name and coordinates (RA, Dec and epoch) are also specified, as well as the observing site, telescope, control unit, filters, date and time of the observation. If external triggering was used, the exposure time also needs to be provided.

The main pipeline script is SHOCpipeline.py, which updates the FITS headers to reflect the information provided by the user, and also reads the appropriate sensitivity and readout noise parameters for the relevant observing mode from an included database. This script also populates three other scripts with the appropriate content for performing the rest of the processing for the particular observation: SHOCscript, PHOTscript and PLOTscript.

SHOCscript slices the data cubes into individual FITS files, with the appropriate timing information for each frame, as well as the calculated airmass, JD (Julian Date) and HJD (Heliocentric Julian Date). The script then creates master bias and master flat frames, and applies them to the target frames.

6.1.5 Aperture photometry

For each of the targets, a number of comparison stars for differential photometry (see §6.1.6) were selected in the same field of view as the target. The number of comparisons were 4 for CAL 83, 4 for RXJ0513 and 2 for MR Vel (see Fig. 6.1). The same comparison stars were used for all the observations. The SIMBAD Astronomical Database⁵ (Wenger et al., 2000) was used to check the fields of the three targets,

⁴<http://www.sao.ac.za/~marissa/SHOCpipeline/>

⁵<http://simbad.u-strasbg.fr/simbad/>

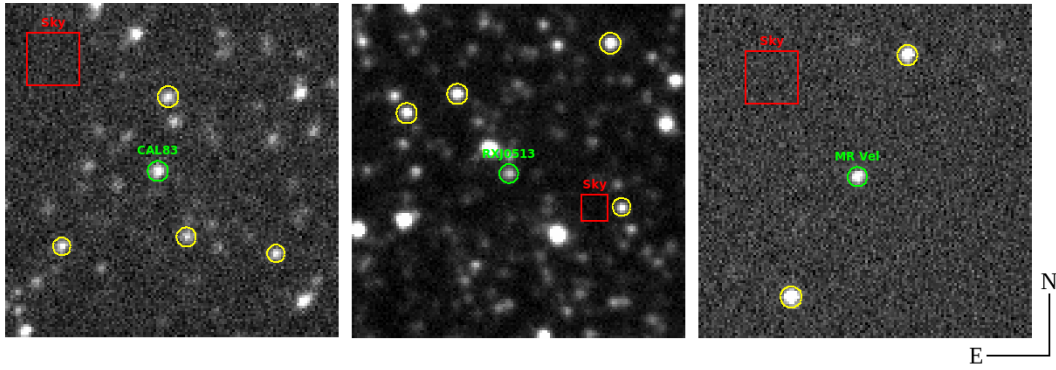


Fig. 6.1: Example images of CAL 83 (left), RXJ0513 (middle) and MR Vel (right) as obtained with SHOC2 on 13 Apr 2013. The target is indicated by the green aperture in the middle of each image, and the comparison stars by the yellow apertures. The aperture sizes represent the mean optimal aperture size (according to the curve-of-growth model) for the respective stars for these particular observations. The red box represents the defined sky region. The field of view of each image is $1.29 \text{ arcmin} \times 1.29 \text{ arcmin}$.

and stars that are known to be variable were not selected as comparisons. The collection of comparison stars used for a certain field is often referred to as an *ensemble*.

After performing the necessary reductions, the positions of the target and comparisons for the particular observations were provided as input to PHOTscript. This script performed aperture photometry and determined the instrumental magnitudes for the target and comparisons by using the **phot** task and related parameter files from the IRAF package **digiphot.daophot**.

The default background region employed by PHOTscript is an annulus around the target aperture. However, the fields containing CAL 83 and RXJ0513 are very crowded, as can be seen in Fig. 6.1, and a sky annulus surrounding the target or comparison aperture may overlap with a nearby target. Even though the field around MR Vel is not crowded, there is another star close by (not visible in Fig. 6.1, but visible in longer exposures) that would be included in an annulus around the target.

Therefore, a single “sky box” was defined for each image in a region that did not contain visible point sources. A $12 \text{ arcsec} \times 12 \text{ arcsec}$ box was used for CAL 83 and MR Vel, and a $6 \text{ arcsec} \times 6 \text{ arcsec}$ box for RXJ0513. The SHOC pipeline was slightly modified to rather determine the mean sky background level from the sky box, and this value was assumed to be representative of the background level throughout the $1.29 \text{ arcmin} \times 1.29 \text{ arcmin}$ image. The same sky background level was thus applied to the target and all the comparisons on the image, but because the sky box level was evaluated for every exposure, variability in the background was still accounted for.

A list consisting of a wide range of source aperture sizes was provided to the **phot** task, which calculated instrumental magnitudes for each of the aperture sizes. Consequently, PHOTscript employed the task **mkapfile** from the IRAF package **digiphot.photcal** to perform aperture-corrected photometry. This task makes use of a curve-of-growth (COG) method (Stetson, 1990) to determine the smallest aperture that maximizes the signal-to-noise (i.e. the turn-over point in the COG). It then extracts the magnitude for each star corresponding to its optimal aperture from the magnitude files created by **phot**, and corrects this magnitude to the value it would have had if the largest aperture in the list had been used. For all the CAL 83 and RXJ0513 observations, the radius of the largest aperture in the provided list was 3.6 arcsec, and for MR Vel it was 4.3 arcsec.

After the computation of the instrumental magnitudes with **phot**, PLOTscript was used to extract lightcurves (i.e. datasets of magnitude versus time) for the target and all the comparisons. In a few of the data frames, the S/N of one or more of the faint stars (target or comparisons) was too low to determine an optimal aperture. For such frames, the measurements of the other stars were also discarded, leaving a set of “filtered” lightcurves for the target and comparisons that were comprised only of measurements from frames where “good” measurements were obtained for the target *and all the comparisons*. The lightcurve time stamps (corresponding to the mid-times of the respective exposures) were converted from UTC to BJD_{TDB}, i.e. with the solar system barycentre as reference point, in the Barycentric Dynamical Time standard. This was accomplished by utilizing the “UTC2BJD” Time Utility of the Ohio State University⁶ (Eastman et al., 2010), which applied the necessary corrections as explained in §4.1.4, and is accurate to a few milliseconds.

It is acknowledged that the best photometric technique for analysing crowded fields involves point spread function (PSF) fitting of the stellar profiles, especially when undertaking absolute photometry. For the current analysis, which is focussed only on obtaining differential photometry measurements to study variability, the method described above is considered to be relatively robust. However, the more involved PSF fitting technique may be utilized in future to refine the quality of the SHOC lightcurves.

6.1.6 Differential photometry

Differential photometry is a process during which the magnitude of the target star in each frame is determined relative to the magnitude of a reference or “comparison” star, usually (and preferably) on the same field as the target (e.g. Howell,

⁶<http://astroutils.astronomy.ohio-state.edu/time>

2006, pp. 121-123). It is assumed that the chosen comparison star does not exhibit any intrinsic variability on the time-scale of the observation, and that any variability in the observed magnitude of the comparison is due to variability in the atmosphere (i.e. airmass and transparency), with possible contributions from any time-dependent properties of the sensitivity of the telescope or instrument. All these effects will affect the target and the comparison on the same field to the same extent.

Assuming the intrinsic luminosity of the comparison to be constant, one can determine a differential correction that would adjust the comparison magnitude to a predefined constant level. Applying the same correction to the target magnitude therefore removes the effects of atmospheric and instrumental variability from the target lightcurve, leaving only its intrinsic variability. If no prior knowledge is available regarding the selected comparison star, there is evidently some risk that the comparison itself may be variable to some extent. In order to minimize this risk, general practice is to select several comparison stars from the same frame, and then apply an averaged differential correction.

Differential photometry enables one to study the variability in the target to a great level of precision, although the absolute flux of the target is not determined. To achieve the latter, an absolute correction should be obtained by observing known standard stars, a process that is known as *absolute photometry*.

As can be seen in Table 6.1, 11 observations of CAL 83 were obtained, 3 of RXJ0513 and 8 of MR Vel. To enable direct comparison of the corrected lightcurves of a particular target, the different lightcurves for the particular target should be calibrated to the same differential “reference system” as far as practically possible. Therefore, after the extraction of filtered instrumental magnitude lightcurves as described in the previous section, the 11 target lightcurves of CAL 83 were joined together to yield a single master target lightcurve. A similar combination of the comparison lightcurves yielded one master lightcurve for each of the 4 comparisons.

A differential correction was then applied to the master target lightcurve by using the 4 master comparison lightcurves. The differential photometry corrections were performed with a Python script written by the author, based on a method similar to the one described by e.g. Everett and Howell (2001). The technical details pertaining to this method are described in Appendix E. The lightcurves associated with RXJ0513 and MR Vel were combined and corrected in a similar manner as for CAL 83.

After the differential correction of the master lightcurves, the corrected lightcurves were again split into the separate observations as listed in Table 6.1. Where multiple

lightcurves of the same target were obtained during a single night, these lightcurves were then joined together. In addition to the 11 individual observations of CAL 83, a single lightcurve for 12 Apr 2013 was then created from observations 03 and 04, as well as a single lightcurve for 16 December 2014 consisting of observations 08, 09 and 10. Similarly, for MR Vel, observations 01, 02 and 03 were combined for 11 Apr 2013, 04 and 05 were combined for 13 Apr and 06 and 07 combined for 14 Apr. These lightcurves will be presented in the next section.

Results

The SHOC lightcurves of CAL 83, RXJ0513 and MR Vel, with the differential corrections applied, are presented in Fig. 6.2, 6.3 and 6.4 respectively. All the lightcurves are presented with a bin size of 50 s, except for the shortest observations of CAL 83 (on 10 and 14 Apr 2013), where a bin size of 10 s was used, and the long CAL 83 observation of 16 December 2014, with a bin size of 100 s. Because no periodic behaviour with periods below ~ 500 s was detected (see §6.1.7), only the longer time-scale variability is illustrated here. To maintain consistency with the notation used in the differential photometry description in Appendix E, the corrected target magnitudes on the vertical axis are indicated by the symbol m_{*0} .

When inspecting the lightcurves, one can see that significant, relatively smooth variations are present in most of them. In the CAL 83 lightcurves, “hump”-like structures are especially apparent, i.e. a smooth rise and decline lasting for $\gtrsim 1000$ s, with an amplitude of up to ~ 0.1 mag. These structures can be seen to commence at ~ 1000 s after T_0 in the 11 Apr lightcurve, at ~ 200 s and again at ~ 6500 s in the 12 Apr lightcurve, at ~ 2700 s on 13 Apr, and perhaps at the start of the 15 Apr observation.

In the longest CAL 83 observation, obtained on 16 December 2014, several of these humps are discernible, commencing at the following times after the observation commences (T_0): ~ 2000 s, ~ 9500 s, ~ 11500 s, ~ 14000 s, and ~ 17500 s. Similar structures with slightly smaller amplitudes and periods also seem to be present in some of the lightcurves. Although these rise-and-decline patterns seem to repeat themselves, they are not strictly periodic; they can be described as quasi-periodic at most. In the 16 Dec 2014 lightcurve, a slower rise and decline starting at $T_0 \sim 6000$ s and lasting for approximately 10000 s, is also evident. No deviation of more than ~ 0.1 mag from the mean was observed within a single night. The maximum change in the mean magnitude from one night to the next is also only ~ 0.1 mag. These modulations in the SHOC lightcurves are of a similar time-scale and amplitude as the modulations observed in the Optical Monitor magnitudes (§4.3.2).

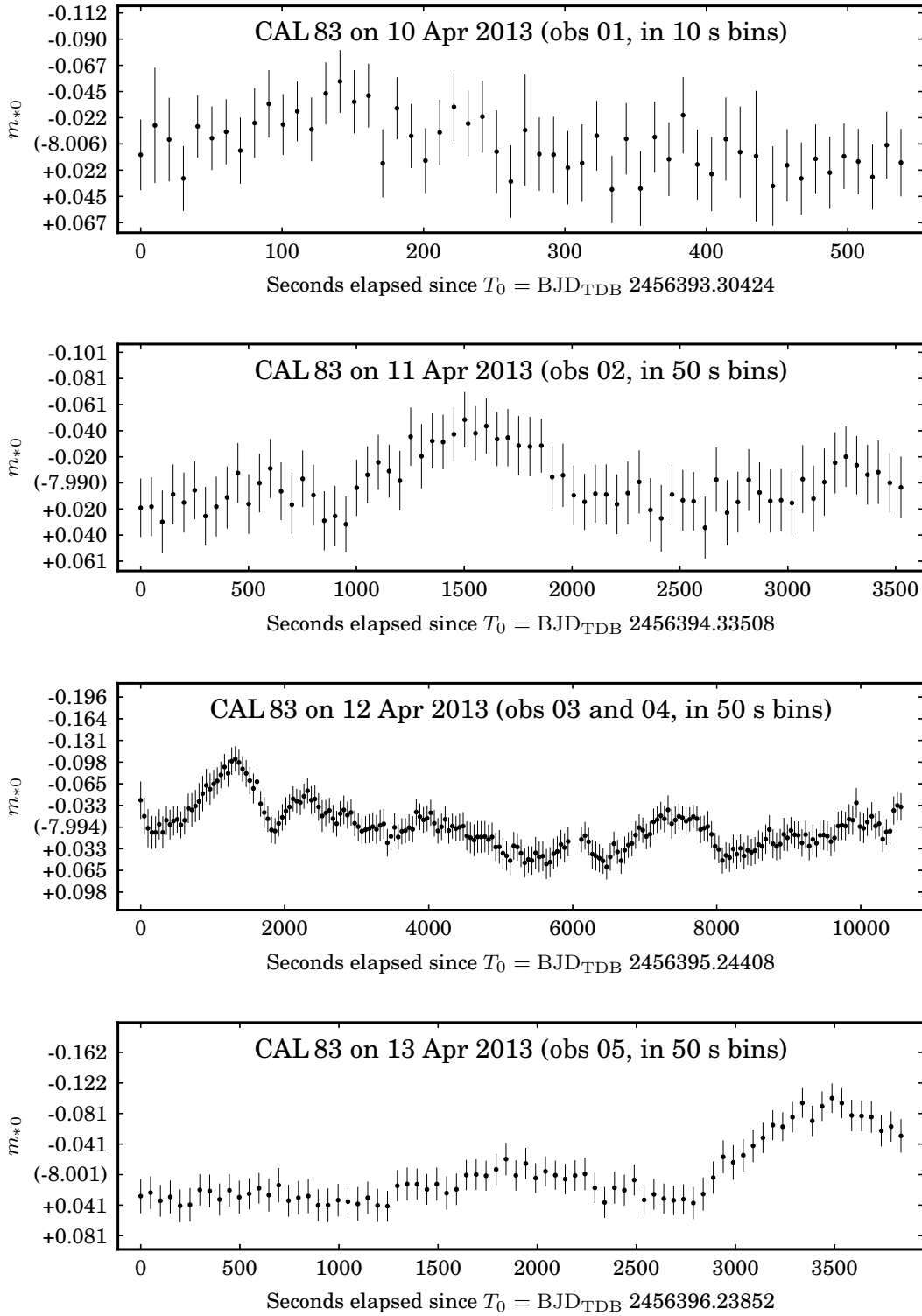


Fig. 6.2: The SHOC lightcurves of CAL 83 after the differential correction. The horizontal axis indicates the seconds elapsed since the first measurement in the observation. On the vertical axis, the mean corrected instrumental magnitude is given in parentheses, with the other markings indicating integer multiples of the standard deviation of the lightcurve, referred to the mean value. The observation number in the respective labels refers to the number allocated in Table 6.1.

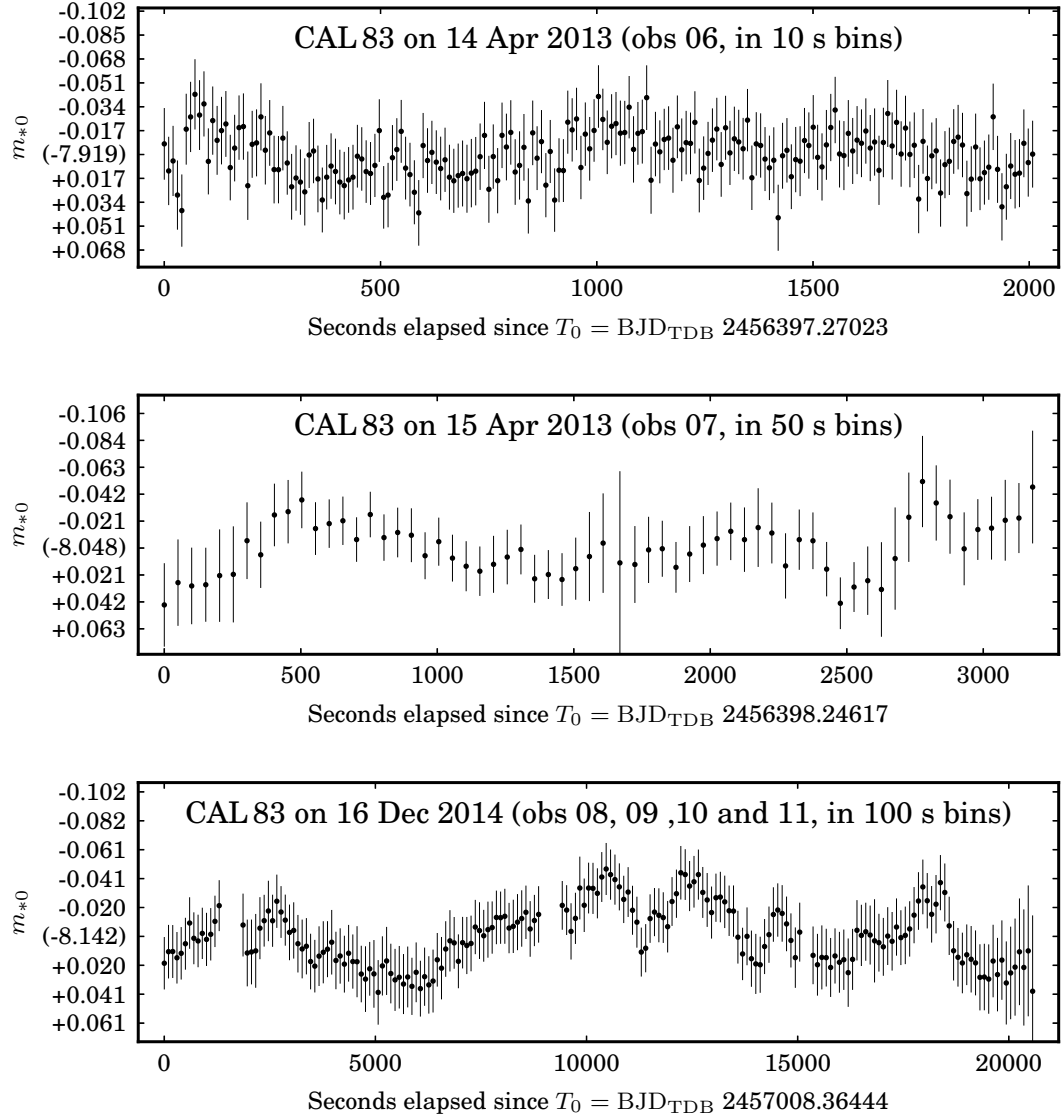


Fig. 6.2: (continued) The SHOC lightcurves of CAL 83 after the differential correction. The horizontal axis indicates the seconds elapsed since the first measurement in the observation. On the vertical axis, the mean corrected instrumental magnitude is given in parentheses, with the other markings indicating integer multiples of the standard deviation of the lightcurve, referred to the mean value. The observation number in the respective labels refers to the number allocated in Table 6.1.

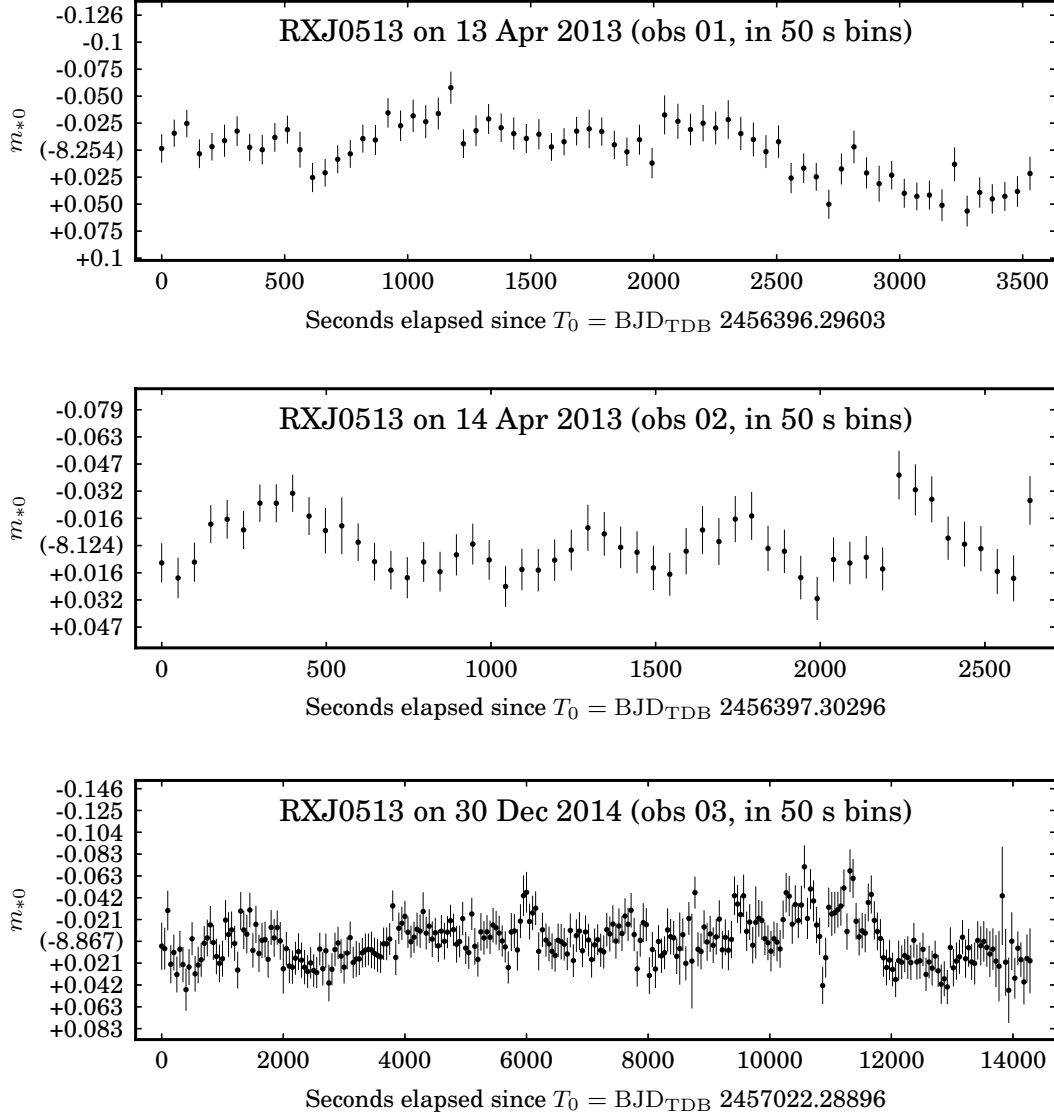


Fig. 6.3: The SHOC lightcurves of RXJ0513 after the differential correction. The horizontal axis indicates the seconds elapsed since the first measurement in the observation. On the vertical axis, the mean corrected instrumental magnitude is given in parentheses, with the other markings indicating integer multiples of the standard deviation of the lightcurve, referred to the mean value. The observation number in the respective labels refers to the number allocated in Table 6.1.

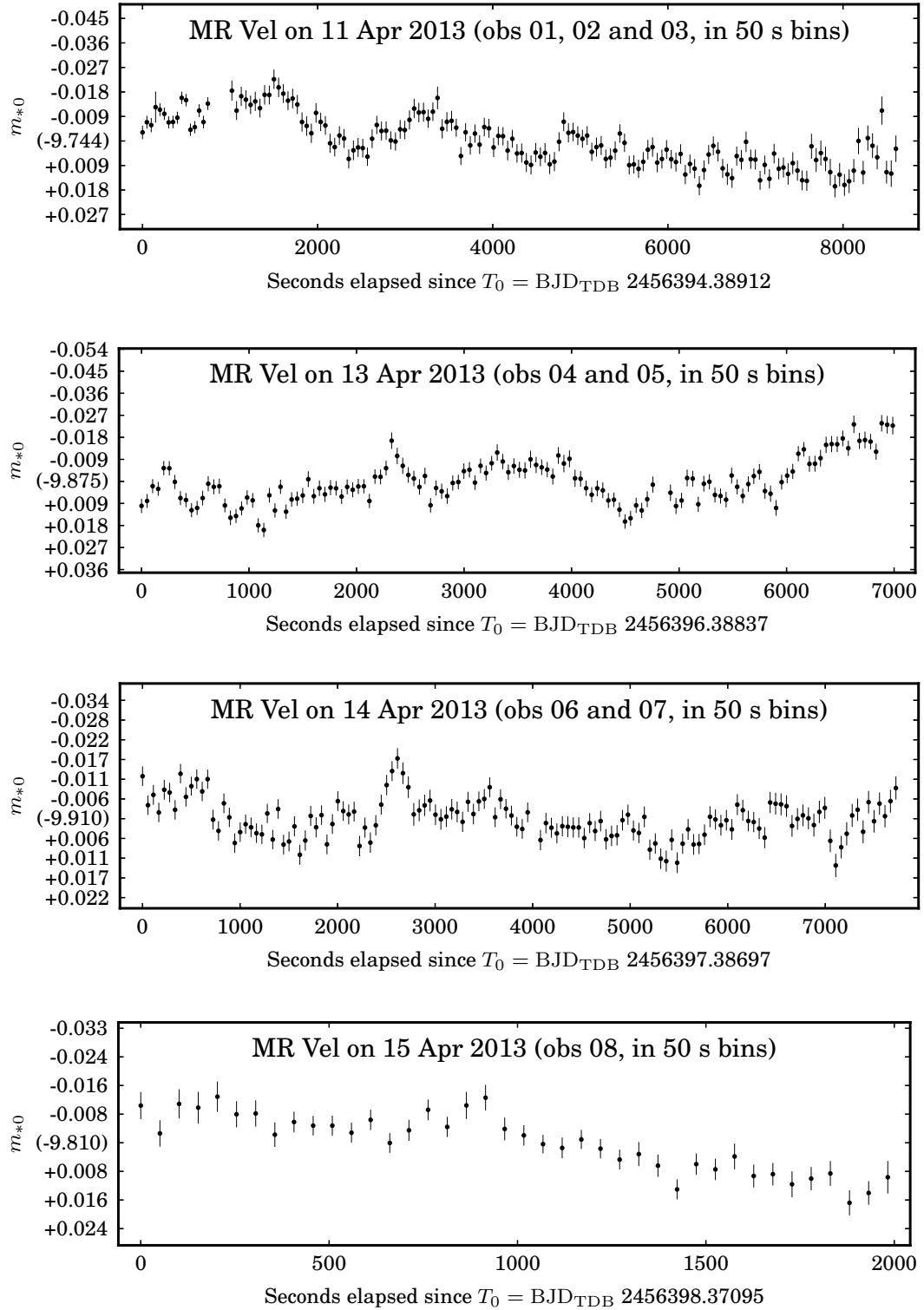


Fig. 6.4: The SHOC lightcurves of MR Vel after the differential correction. The horizontal axis indicates the seconds elapsed since the first measurement in the observation. On the vertical axis, the mean corrected instrumental magnitude is given in parentheses, with the other markings indicating integer multiples of the standard deviation of the lightcurve, referred to the mean value. The observation number in the respective labels refers to the number allocated in Table 6.1.

The lightcurves of RXJ0513 also exhibit variability with a full amplitude of ~ 0.1 mag. On 14 Apr 2013, the variability has a smooth appearance, similar to some of the CAL 83 lightcurves, with the appearance of quasi-periodic oscillations. In the RXJ0513 lightcurves of 13 Apr 2013 and 30 Dec 2014, the variability seems to be more erratic, and in the 30 Dec 2014 lightcurve “flare”-like brightness increases are superimposed on modulations on longer time-scales. The source was brighter in Dec 2014 than in Apr 2013 by ~ 0.7 mag. This is consistent with the OGLE-IV lightcurve (Fig. 6.7), according to which the source was in an optical low state in Apr 2013, and in an optical high state in Dec 2014.

The MR Vel lightcurves also exhibit apparent quasi-periodic modulations, but with a full amplitude of only $\lesssim 0.05$ mag. Flare-like structures were observed at times of ~ 2200 s after T_0 on 13 Apr, and ~ 2600 s after T_0 on 14 Apr.

Discussion

The orbital periods of the three sources under consideration are: 1.047529 d for CAL 83 (semi-amplitude ~ 0.1 mag), 0.762956 d for RXJ0513 (semi-amplitude ~ 0.03 mag, Rajoelimanana et al., 2013), and 4.0287822 d for MR Vel (semi-amplitude ~ 0.2 mag, Schmidtke and Cowley, 2001). Although the orbital modulation in these systems should be present in the SHOC lightcurves, the timespan of even the longest of the SHOC observations are still significantly shorter than the orbital periods, therefore it is evident that variability on shorter time-scales is superimposed on the presumed underlying orbital modulation.

These modulations are expected to originate in the accretion disc, which is the dominant source of optical emission in SSSs. For a semi-quantitative evaluation of the possible emission regions associated with the optical modulations, it is useful to evaluate the dimensions of the accretion disc and also the corresponding Keplerian periods associated with the disc in these systems. These estimates are summarized below.

The binary parameters of CAL 83 have been discussed in §5.6.2. Just like for CAL 83, there is still considerable uncertainty in the binary parameters of RXJ0513 and MR Vel. Although the orbital periods of these sources are well known, and the WD mass can be constrained from X-ray model fits, it is uncertain whether the radial velocity semi-amplitude K_1 of the WD is well represented by the emission lines arising mainly in the asymmetric accretion disc and winds in the system. As the secondary stars in these systems are not observable in the optical spectra and their masses can thus not be determined directly, unambiguous constraints on the binary inclination and secondary mass are very difficult to achieve.

However, by imposing the requirement of a shrinking secondary Roche lobe, which is required to drive mass transfer, semi-quantitative constraints can nevertheless be obtained for binary parameters. It has been shown in §5.6.2 that for a $\sim 1.3 M_{\odot}$ WD, a lower limit to the secondary mass which would cause the secondary Roche lobe to shrink upon mass transfer is $M_2 > 0.79 M_{\text{wd}} \sim 1.0 M_{\odot}$.

Adopting $P_{\text{orb}} = 0.762956$ d for RXJ0513, and the lower limit of $M_2 > 1.0 M_{\odot}$, yields an orbital separation of $a \sim 3.2 \times 10^{11}$ cm according to Eq. (2.1). The corresponding Roche lobe radii are $R_{\text{L},1} \sim 1.3 \times 10^{11}$ cm and $R_{\text{L},2} \sim 1.2 \times 10^{11}$ cm (Eq. (2.7)). The circularization radius according to Eq. (2.31) is $\sim 4.4 \times 10^{10}$ cm. Assuming tidal truncation of the disc at $0.7 R_{\text{L},1}$ (Papaloizou and Pringle, 1977), the outer disc limit is at $R_{\text{out}} \sim 9.1 \times 10^{10}$ cm. The Keplerian period associated with this value of R_{out} is ~ 13000 s.

The $\sim 7 \times 10^{10}$ cm radius of a $1 M_{\odot}$ main sequence star is much smaller than the radius of its Roche lobe. For Roche lobe overflow to take place, the donor would need to be an evolved star. The presumably convective envelope of such a secondary star would expand upon the loss of its outer layers through mass transfer, which can in turn aid in driving the Roche lobe overflow (e.g. Meintjes, 2002 and references therein).

The radial velocity semi-amplitude K_1 inferred from the optical emission lines of RXJ0513 is $\lesssim 15$ km s $^{-1}$ (§3.2.3). Adopting these parameters, together with the lower limit of $M_2 > 1.0 M_{\odot}$, an upper limit of $i \lesssim 6^{\circ}$ is obtained for the binary inclination (a similar limit of $i \lesssim 7^{\circ}$ was calculated by Southwell et al., 1996a).

The orbital inclination of MR Vel should not be too high, as no eclipses are observed, and the high outflow velocities observed in the optical spectra are comparable to the WD escape velocity, i.e. the angle between the line of sight and a jet perpendicular to the orbital plane can not be too large. According to Schmidtke et al. (2000), the inclination can not be too low either, as the large implied secondary mass and luminosity would then make it visible in the optical spectrum, which it is not. Schmidtke et al. proposed $i \lesssim 65^{\circ}$ for MR Vel.

Although the component masses are not well constrained, various considerations have shown that the WD mass is just below the Chandrasekhar limit (see §3.3). Applying $P_{\text{orb}} = 4.0287822$ d, an upper limit of $i \lesssim 65^{\circ}$ for the inclination and $M_{\text{wd}} \sim 1.4 M_{\odot}$, together with $K_1 \sim 84$ km s $^{-1}$ (Motch, 1996), yields a relatively large mass function and $M_2 \sim 1.4 M_{\odot}$. Smaller inclinations would yield larger values for M_2 . This value obeys the condition of $M_2 > 0.79 M_{\text{wd}}$ discussed before, and is

close to the estimations of other authors, who proposed the secondary to be a giant with mass $\gtrsim 1 M_{\odot}$. The associated binary separation is $a \sim 1.0 \times 10^{12}$ cm, with Roche lobe radii of $R_{L,1} \sim 4.0 \times 10^{11}$ cm and $R_{L,2} \sim 3.9 \times 10^{11}$ cm, and a circularization radius of $R_{\text{circ}} \sim 1.3 \times 10^{11}$ cm. Estimating the outer disc radius to be at $0.7R_{L,1}$, yields $R_{\text{out}} \sim 2.8 \times 10^{11}$ cm, and an associated Keplerian period of ~ 67000 s.

Quasi-periodic oscillations with periods of ~ 5.5 , ~ 6.2 and ~ 7.4 h have been reported previously in MR Vel. It is noted that the longest of these periods are still shorter than the ~ 67000 s Keplerian period expected to be associated with the outer edge of the disc. However, the SHOC observations were not long enough to investigate whether these long periods could still be detected.

In §5.6.2, the outer radius of the accretion disc in CAL 83 was estimated to be $R_{\text{out}} \sim 1.1 \times 10^{11}$ cm, which is associated with a Keplerian period of ~ 18000 s. If the magnetic field of the WD is weak, the disc can extend to the surface layers of the WD, and the inner disc radius can be approximated with $R_{\text{in}} \sim 10^9$ cm, corresponding to a Keplerian period of ~ 15 s for a massive WD.

The Keplerian periods calculated above serve to show that the variability on time-scales of ~ 1000 s observed in the SHOC lightcurves of the three sources may be associated with Keplerian periods in the disc closer to the accreting WD, possibly of turbulent eddies that dissipate after one or two orbits. However, the extended rim of the accretion disc is expected to be the dominant source of (reprocessed) optical emission from the system. (Schandl et al., 1997; Meyer-Hofmeister et al., 1997). Therefore, the optical modulations may also be related to small changes in the structure of the rim itself, or perhaps the movement of blobs inside the rim.

It has been proposed (§4.7.4) that the ~ 67 s X-ray periodicity in CAL 83 can be explained in a similar way as the DNOs (dwarf nova oscillations). As explained in §2.7.4, QPOs (quasi-periodic oscillations) occurring simultaneously with DNOs often exhibit the period relation $P_{\text{QPO}}/P_{\text{DNO}} \sim 15$. Adopting $P_{\text{DNO}} \sim 67$ s leads to a period of $P_{\text{QPO}} \sim 15 \times 67 \text{ s} \sim 1000$ s for a potential DNO-related QPO in the system. It is interesting to note that this predicted time-scale is of the same order as the quasi-periodic modulations that have been observed in the SHOC lightcurves.

6.1.7 Lomb-Scargle analysis

All the SHOC lightcurves for the three sources were consequently subjected to a Lomb-Scargle (LS) timing analysis procedure with the Starlink PERIOD package, as described in §4.4.1. As before, the significance levels over Gaussian white noise were determined with Monte Carlo simulations.

Initially, the individual lightcurves were detrended by subtracting a second-order polynomial, and LS periodograms were created for these lightcurves, using the original bin sizes corresponding to their respective exposure times. The maximum frequency (shortest period) investigated corresponds to 2 times the value of the exposure time as given in Table 6.1. The shortest periods that could thus be investigated were 1 s for CAL 83 (observation 06), 3 s for RXJ0513 (observation 02), and 0.1116 s for MR Vel (observation 07). All of these values are shorter than the ~ 4 s break-up period of a $\sim 1.3 M_{\odot}$ (refer back to §4.7.2). The Lomb-Scargle analysis of these lightcurves did not reveal any periodicities in these sources above a 99.73% level at frequencies above 2 mHz, i.e. periods shorter than 500 s.

In particular, no trace of the ~ 67 s period was found in the SHOC lightcurves of CAL 83. Referring to the OGLE-IV lightcurve in Fig. 6.6 in §6.2, one can see that the Apr 2013 observations (centred on $\text{BJD}_{\text{TDB}} \sim 2456395$), as well as the observations on the night of 16 December 2014 ($\text{BJD}_{\text{TDB}} \sim 2457008$), took place when CAL 83 was in an optical low state. Unfortunately, no simultaneous X-ray observations are available, but the anti-correlation model implies that the source was in an X-ray high state during this time.

As shown in §4.4, the ~ 67 s X-ray periodicity was detectable during all the *XMM-Newton* observations during an X-ray high state. It is therefore very likely that the ~ 67 s periodicity was present in *X-rays* during the SHOC observations. However, its $\lesssim 10\%$ amplitude in the X-ray waveband is probably too small to be visible in the optical at a detectable level after scattering from the disc.

The question now remains as to whether the ~ 67 s periodicity may be visible in the optical during optical high, X-ray low states. As the WD envelope expands during the optical high state due to enhanced accretion (§3.1.1), it is possible that the equatorial belt proposed as the origin of the ~ 67 s periodicity (§4.7.4) shifts to a somewhat larger radius with a lower effective temperature (and a slightly longer period). This might cause the periodicity to be detectable in the optical waveband, but this possibility remains to be investigated in follow-up studies.

The MR Vel observations of 14 Apr 2013 did exhibit a very sharp peak at ~ 11 mHz, but this peak was present not only in the target, but also in both the comparisons, in all the lightcurves before and after the differential correction. It is therefore concluded that this periodicity is not inherent to either one of the stars, but arose from some periodic local illumination of the detector.

Consequently, the Lomb-Scargle analysis was focussed on frequencies below 2 mHz,

i.e. periods longer than 500 s. A lightcurve binning of 50 s was used, and the lightcurves were now detrended by subtracting a first-order polynomial. These periodograms are shown in Fig. 6.5. Several peaks were present above the 99.73% significance level obtained from the white noise simulations, and many of them may be associated with the quasi-periodic modulations observed in the lightcurves. However, several of the periodograms exhibit a red noise power-law character, which complicates the calculation of the significance of these peaks. The estimation of the true significances of these peaks with respect to the red noise will form part of a more detailed follow-up study.

6.2 Photometry with MACHO and OGLE

CAL 83 has been monitored as part of the MACHO (MAssive Compact Halo Objects) Project from January 1993 up to January 2000. MACHO monitoring observations of RXJ0513 extended from July 1992 to January 2000. These observations were performed with the 1.27-m telescope at Mount Stromlo Observatory in Australia, and have been published by various authors, as discussed in Chapter 3. During the MACHO observations, measurements in “red” ($\sim 6300\text{--}7600\text{ \AA}$) and “blue” ($\sim 4500\text{--}6300\text{ \AA}$) bands were obtained by making use of a dichroic beamsplitter and filters (Alcock et al., 1996).

Since September 2001, regular observations of CAL 83 and RXJ0513 have also been obtained as part of the OGLE (Optical Gravitational Lensing Experiment) Project (Rajoelimanana et al., 2013), using the 1.3-m Warsaw telescope at Las Campanas Observatory. Phase III extended up to Apr 2009, while phase IV started in March 2010, and is still ongoing. The data for these sources are available through the X-Ray variables OGLE Monitoring (XROM) website⁷ (Udalski, 2008). The OGLE-III lightcurves have been discussed together with the MACHO data by Rajoelimanana et al. (2013), while the OGLE-IV lightcurves have to our knowledge not yet been published. In this section, the superorbital variability of CAL 83 and RXJ0513 will be characterized by making use of the MACHO and OGLE-III *and* -IV lightcurves, as an extension of the work published by Rajoelimanana et al. (2013).

The MACHO and OGLE lightcurves of CAL 83 and RXJ0513 are provided in Fig. 6.6 and Fig. 6.7. The mid-exposure time stamps of the observations were converted to the BJD_{TDB} time standard with the Time Utilities of the Ohio State University⁸ (Eastman et al., 2010). For CAL 83, more measurements were available in the MACHO blue filter than in the red, therefore the blue filter magnitudes are shown. In the RXJ0513 MACHO data, the red filter measurements were in the majority, and

⁷<http://ogle.astrouw.edu.pl/ogle4/xrom/xrom.html>

⁸<http://astrutils.astronomy.ohio-state.edu/time>

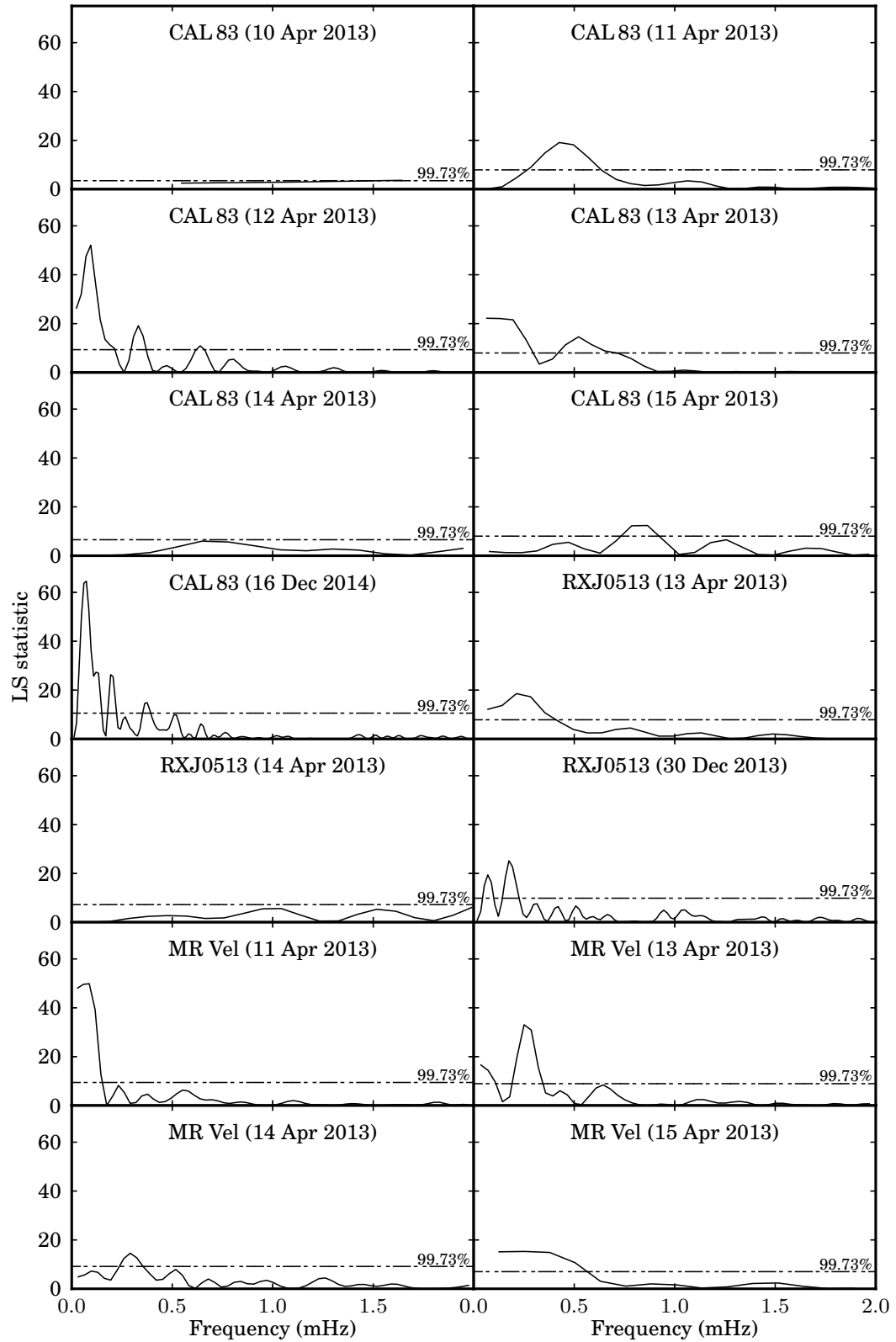


Fig. 6.5: Lomb-Scargle periodograms of the SHOC observations of CAL 83, RXJ0513 and MR Vel, in the 0-2 mHz range. The lightcurve binning is 50 s. The 99.73% significance level over white noise is indicated.

are therefore shown rather than those in the blue filter. The quasi-periodic transitions between optical high and low states over the last 20+ years are evident in the lightcurves.

Consequently, a Lomb-Scargle analysis aimed at characterizing the long-term variability was performed by combining the three lightcurves for each target respectively, focussing on periods >100 d. The bandpass of the MACHO filters are of course somewhat different from the I filter used for the OGLE data, and the values can thus not be compared directly. However, we are not so much interested in the absolute level of the magnitude as in the time-scale of its long-term variability relative to its mean value.

Therefore, in preparation of the Lomb-Scargle analysis, the MACHO, OGLE-III and OGLE-IV lightcurves were detrended by subtracting from each respective lightcurve its mean, and dividing it by its standard deviation. These detrended lightcurves were then joined together, and a Lomb-Scargle periodogram in the $0-0.01 \text{ day}^{-1}$ frequency range was created for each target, as shown in Fig. 6.8. The Lomb-Scargle periodograms exhibit multiple peaks at low frequencies, corresponding to the long-term modulations visible in the lightcurves.

The strongest peak in the CAL 83 periodogram is at 433 ± 12 d, and in the RXJ0513 periodogram the strongest peak is at 168 ± 2 d. These values are in good agreement with the ~ 450 d and ~ 168 d obtained by Rajoelimanana et al. (2013) from only the MACHO and OGLE-III data, and the shapes of the periodograms are also similar to those obtained in the latter paper. Therefore, no dramatic changes in the nature of the cycles have occurred during the last 5 years covered by the OGLE-IV lightcurves.

Although Rajoelimanana et al. (2013) have determined updated values for the orbital periods of CAL 83 and RXJ0513 from the MACHO and OGLE-III lightcurves, the last full orbital ephemeris, i.e. including an accurate time of minimum light, has been published by Schmidtke et al. (2004) and Cowley et al. (2002) respectively for the two sources. A very important aspect of follow-up work would be to determine a full updated orbital ephemeris for both CAL 83 and RXJ0513, using the ~ 23 year baseline provided by the MACHO, OGLE-III and OGLE-IV lightcurves. This would also enable one to calculate accurate orbital phases for the recent and planned spectroscopic observations, so that the possible orbital phase dependence of the spectral properties can be investigated in more detail.

The regular monitoring of CAL 83 and RXJ0513 by MACHO and OGLE has played a key role in our understanding of the accretion modes in these systems. Periodic analyses of these long baseline lightcurves provide the opportunity to constrain the or-

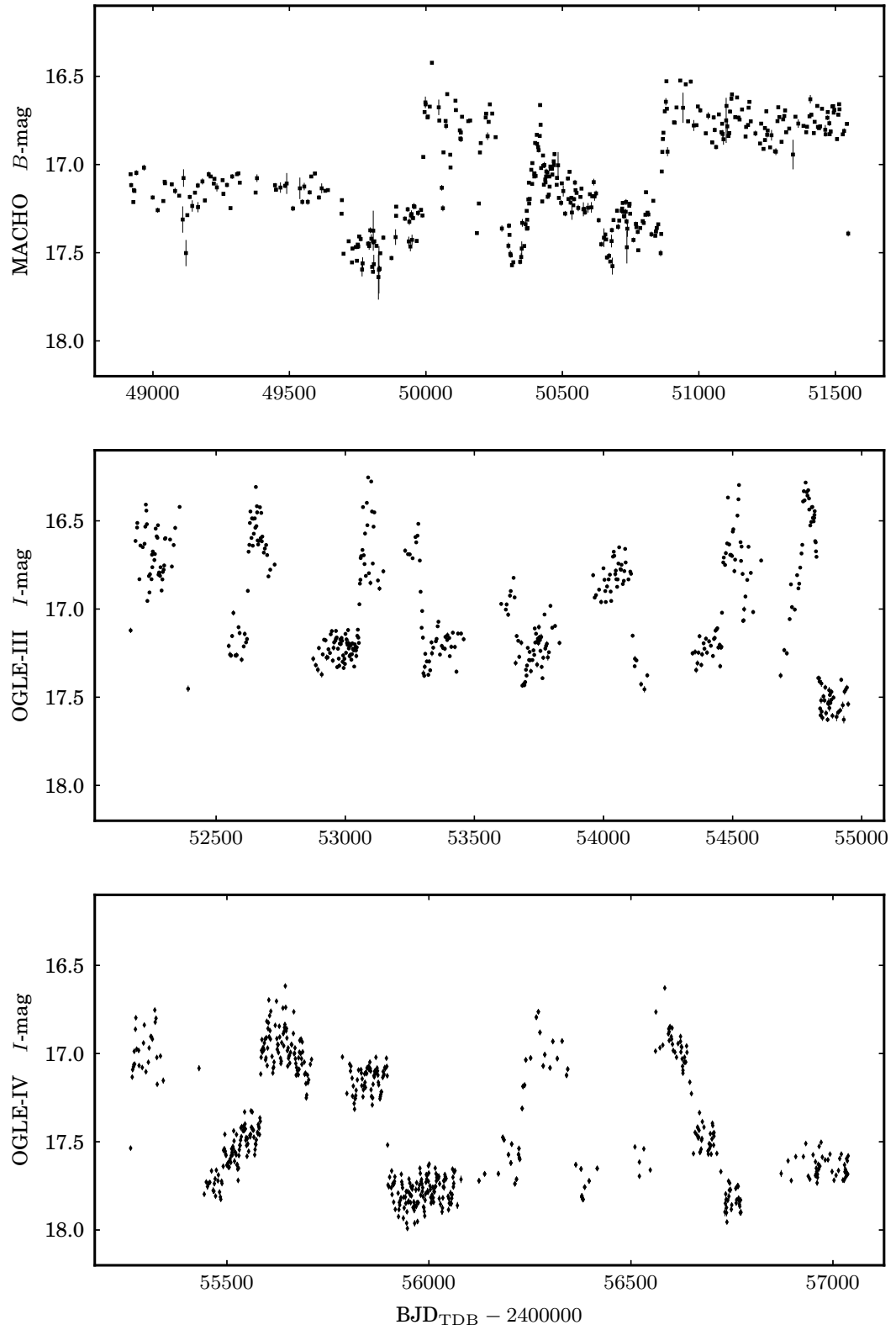


Fig. 6.6: The MACHO (top), OGLE-III (middle) and OGLE-IV (bottom) lightcurves of CAL 83. The cyclic changes of ~ 1 mag between optical high and low states are evident.

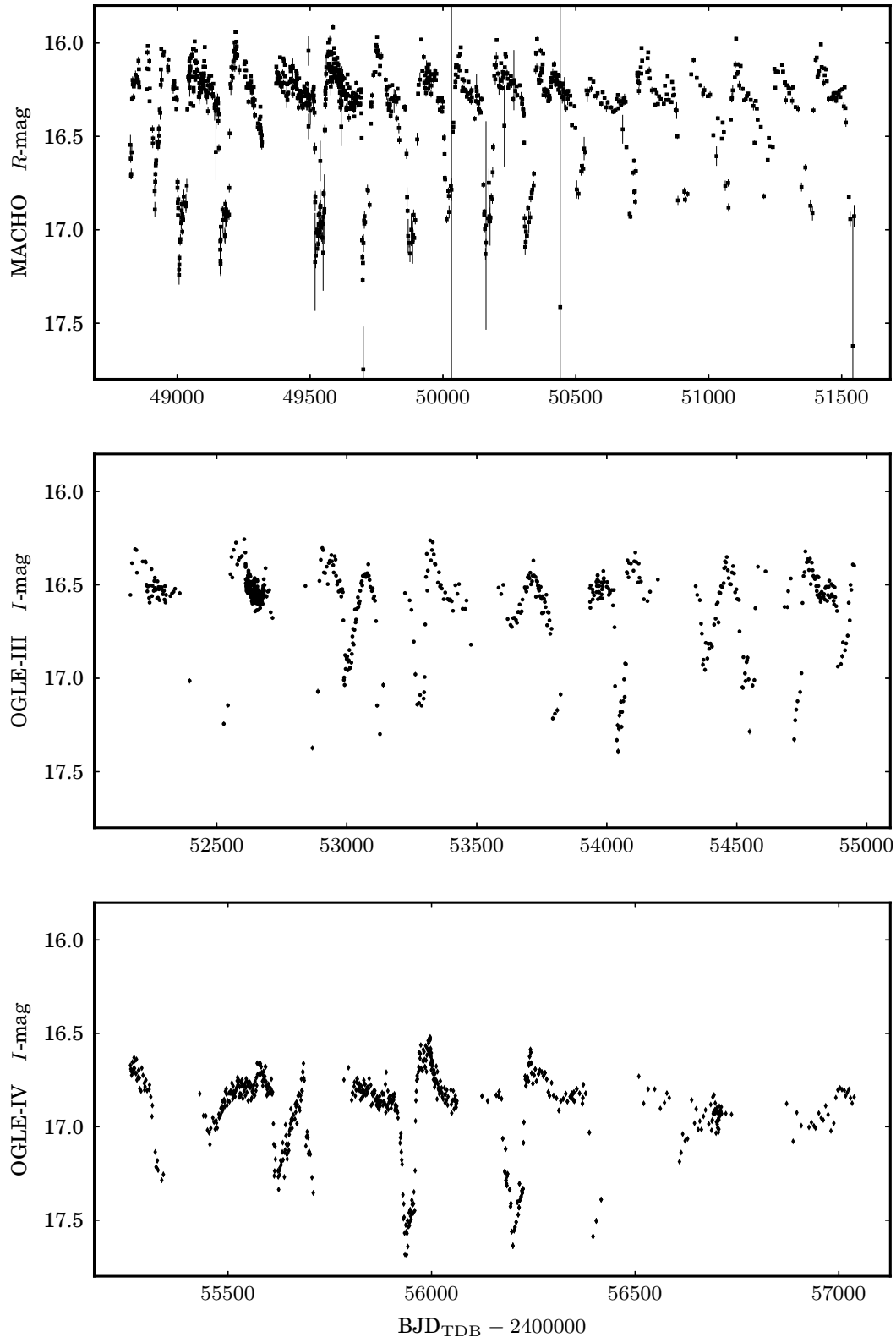


Fig. 6.7: The MACHO (top), OGLE-III (middle) and OGLE-IV (bottom) lightcurves of RXJ0513. The cyclic changes of ~ 1 mag between optical high and low states are evident.

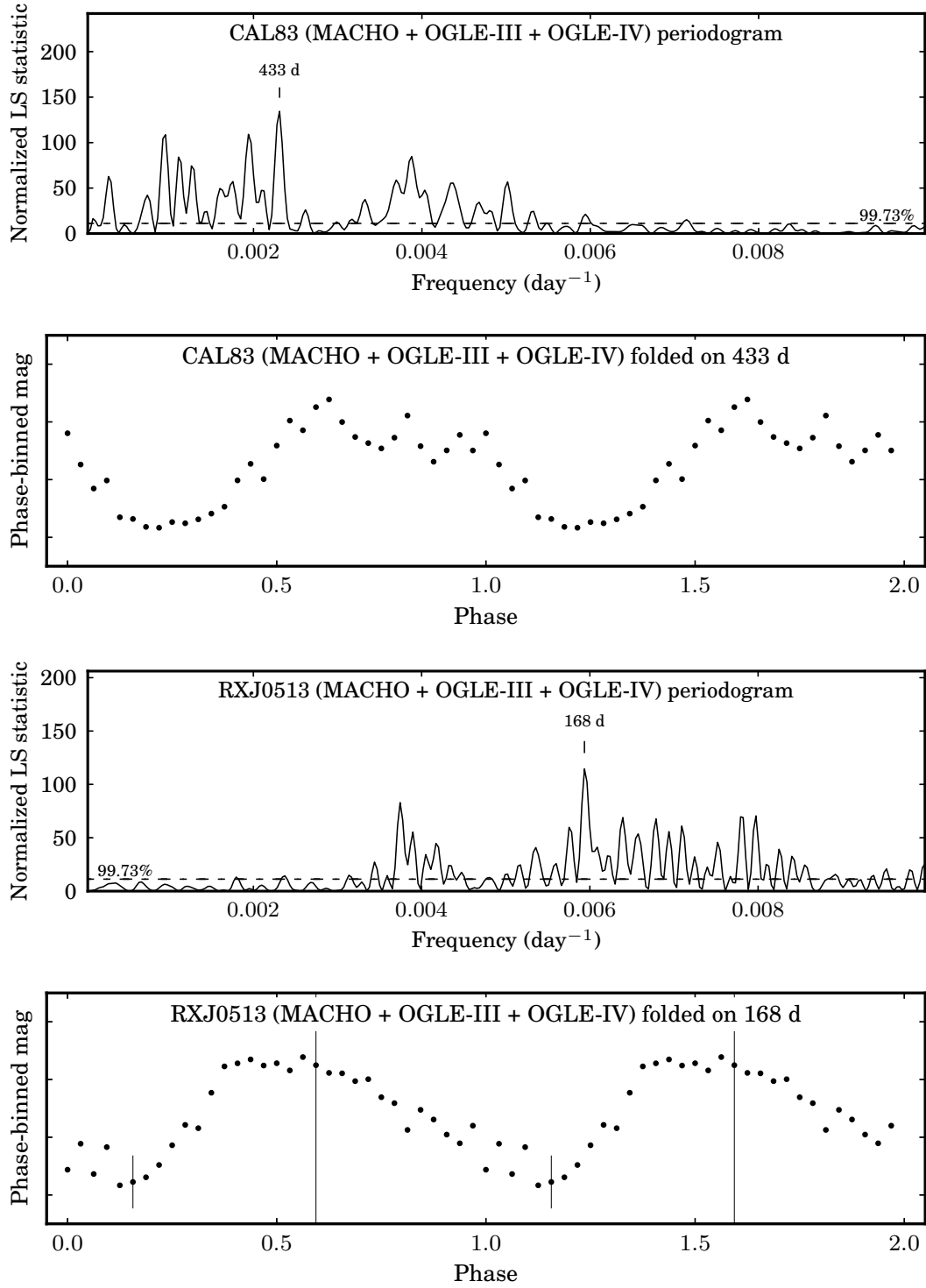


Fig. 6.8: Periodograms and folded lightcurves of CAL 83 and RXJ0513 from the MACHO and OGLE data. Each lightcurve has been folded on the strongest peak in the corresponding periodogram, with the starting point of the MACHO lightcurves as zero point. The 99.73% significance level over white noise is shown. The vertical scale of the folded lightcurves represents phase-binned detrended magnitudes, the values of which are not physically meaningful and are therefore not shown. The full amplitude of the modulations is ~ 1 mag for both sources, as can be seen in Fig. 6.6 and 6.7.

bital and superorbital periods to unprecedented accuracy. However, the lightcurves themselves also provide an immensely useful reference framework in terms of the epochs of optical low and high states, with which other observational results can be compared to better understand the behaviour of these sources in the two distinct optical and X-ray states. Continued OGLE-IV monitoring will be invaluable in future work on CAL 83 and RXJ0513.

Chapter 7

Conclusion

In this study, an investigation of accretion-related variability on various time-scales has been undertaken in the supersoft X-ray binaries CAL 83, RXJ0513 and MR Vel. The main target under investigation was CAL 83, the prototypical source of this class in the LMC. It has been known for quite some time that CAL 83 and RXJ0513 exhibit anti-correlated variability in X-ray and optical flux on superorbital time-scales of several hundred days. Various authors have explained this anti-correlation by limit cycle models involving a contracted white dwarf (WD) envelope with a relatively high effective temperature during the X-ray high, optical low state, and an expanded WD envelope with a lower effective temperature during the X-ray low, optical high state.

The unpublished OGLE-IV lightcurves of CAL 83 and RXJ0513 extending from March 2010 up to 15 January 2015 have been presented. These lightcurves exhibit the same quasi-periodic superorbital modulations as the earlier MACHO and OGLE-III lightcurves. The time-scale of the superorbital cycles in RXJ0513 (~ 168 d) is substantially shorter than in CAL 83 (~ 433 d). The faster progression of the cycles in RXJ0513 is probably due to a higher accretion rate in this source than in CAL 83, a notion which is supported by the slightly brighter optical magnitude of RXJ0513. Follow-up work on the OGLE lightcurves will include the calculation of a fully updated orbital ephemeris for both CAL 83 and RXJ0513.

The archival *XMM-Newton* lightcurves of CAL 83 exhibit significant variability in the X-ray and near-UV wavebands. It has been shown that the average X-ray temperature, as quantified by the hardness ratio of the spectrum, is correlated with the average X-ray flux from observation to observation. This provides confirmation of the temperature-related superorbital cycles mentioned above. A similar correlation has also been observed on short time-scales within the individual observations. It is proposed that the short time-scale increases in the X-ray flux and associated temperature could be related to local pressure enhancements in the WD envelope,

caused by e.g. small increases in the accretion rate, yielding an elevated temperature and X-ray flux.

A detailed period analysis of the *XMM-Newton* lightcurves of CAL 83 has revealed a variable X-ray modulation with a period of ~ 67 s. This periodicity was detected in all the X-ray high state lightcurves obtained during the 2000-2009 time period, and in the X-ray low state lightcurve with the highest count rate. The model proposed to explain the behaviour of this periodicity is related to the low-inertia magnetic accretor (LIMA) model developed by Warner (1995*b*) and Warner and Woudt (2002) to explain the observed properties of dwarf nova oscillations (DNOs).

In this model, the ~ 67 s periodicity originates in a belt-like structure in the envelope at the boundary with the inner accretion disc, which is weakly coupled to the WD core by a WD magnetic field with a strength of $\sim 10^5$ G. Although the LIMA model has been well described for dwarf novae, a detailed theoretical investigation of the properties of a similar model in supersoft X-ray binaries with their much higher accretion rates and more massive envelopes, constitutes a very interesting follow-up project. Periodic analyses of the *XMM-Newton* X-ray lightcurves of RXJ0513 and MR Vel did not reveal any similar, persistent short time-scale periodicities.

The SALT spectra of CAL 83 are dominated by Balmer and He II emission lines, as well as ionized C, N and O emission. As reported by previous authors, the Balmer and He II lines exhibit broad wing structures and P Cyg profiles, of which the velocity dispersion indicates emission from a wind with a large opening angle, presumably originating in the inner regions of the disc. Previously published spectra have shown these wing structures to be blueshifted relative to the main emission line component at some epochs, and redshifted at others. However, all the SALT spectra indicate a blueshifted wing, although the last spectrum is compatible with either a blue- or a redshifted wing.

The velocity displacement of the H α wing relative to the main component in the SALT spectra changed significantly on time-scales of several weeks. This has been ascribed to a precession in the accretion disc and associated outflow by Crampton et al. (1987), who also suggested a potential precession period of ~ 69 d. Although the current SALT results are compatible with this period, obtaining a high S/N spectrum of the source approximately every two weeks over ~ 5 months during an optical high state would enable one to constrain such a potential precession period.

A qualitative comparison of the intensity of the O VI $\lambda 5290$ emission line with respect to the He II $\lambda 5411$ emission line has revealed that the O VI line is of similar strength to the He II $\lambda 5411$ line during optical low, X-ray high states, while O VI $\lambda 5290$ is

barely detectable in the SALT spectra during optical high, X-ray low states. The O VI $\lambda 5290$ line strengths reported by previous publications are in favour of these findings.

The variability of the O VI $\lambda 5290$ line strength serves as evidence for cyclic temperature variations between the optical low and high states. It is also in excellent agreement with the atmospheric expansion/contraction models. The O VI emission is expected to originate in the hot inner regions of the accretion disc very close to the WD. The higher intensity of the O VI line during optical low states is to be expected, as the contracted state of the WD envelope at these times may allow optically thin emission from the inner disc to be observed. The expanded envelope during the optical high state would obscure the hot inner disc and attenuate the observed intensity of O VI emission.

The SALT optical spectra of RXJ0513 are very similar to those of CAL 83, although the Balmer and He II emission lines of this source exhibit sharp bipolar satellite features in addition to wing- and absorption structures, similar to what has been reported in previous publications. These structures indicate the presence of a collimated jet in the system, possibly with an additional outflow component with smaller associated velocities and a larger opening angle. The high-velocity P Cyg absorption structures in the published X-ray spectra of RXJ0513 support the occurrence of outflows from regions close to the WD.

The new SALT spectra of RXJ0513, as well as previously published spectra, indicate that the intensity of the O VI $\lambda 5290$ line is lower than that of the He II $\lambda 5411$ line during optical high states, which is consistent with what was observed in CAL 83. However, spectroscopic observations of RXJ0513 during its optical low state are needed to establish whether the intensity of O VI $\lambda 5290$ emission is indeed anti-correlated with the optical continuum flux, as seems to be the case in CAL 83.

To further investigate this potential anti-correlation and its quantitative implications for the accretion modes, a high S/N spectrum of the spectral region containing the O VI $\lambda 5290$ line, as well as the O VI $\lambda\lambda 3811, 3834$ lines, should be obtained for each of these sources approximately on a monthly basis over a full superorbital cycle. A more extended utilization of the existing SALT datasets is also planned, especially in terms of employing the measured line parameters and ratios to investigate potential kinematic models for the observed outflows.

In contrast to the other two sources, MR Vel seems to be a relatively constant source, and no evidence of dramatic variability in its optical or X-ray flux has yet been observed. The properties of its spectral lines have not changed significantly

over the last >20 years. The $H\alpha$ emission line exhibits a P Cyg profile, as well as a sharp emission feature redshifted by approximately $+4000 \text{ km s}^{-1}$ with respect to the main component. These features, as well as the high-velocity P Cyg absorption in the X-ray spectra reported in previous publications, are indicative of a jet-like outflow in the system.

SHOC observations of CAL 83 during two different optical low states did not reveal the $\sim 67 \text{ s}$ periodicity observed in X-rays, or any other periodic variability at high frequencies, in the optical lightcurves. During optical high states, the expansion of the WD envelope may cause the equatorial belt, according to the LIMA-type model, to shift to slightly larger radii. This would imply a slightly larger period and a lower effective temperature to be associated with the belt. This could imply that the $\sim 67 \text{ s}$ periodicity may be observed in the optical during optical high states, which is an important aspect to investigate with follow-up SHOC observations.

However, the SHOC lightcurves of CAL 83 do exhibit significant quasi-periodic variability on longer time-scales of $\gtrsim 1000 \text{ s}$. If the $\sim 67 \text{ s}$ periodicity represents the equivalent of a DNO, the predicted period of a DNO-related QPO in the system would be $P_{\text{QPO}} \sim 15 \times 67 \text{ s} \sim 1000 \text{ s}$. It is interesting that this value is of the same order as the quasi-periodic variability that has been observed, and this relation should be explored in more detail as part of the theoretical investigations into the nature of the modified LIMA model.

The SHOC lightcurves of RXJ0513 and MR Vel also exhibited variability on similar time-scales, as well as a few flare-like brightness increases. These phenomena in CVs are usually associated with (MHD) turbulent activity. Similarly, in these SSSs, the detected optical modulations can be considered as phenomenological evidence of the presence of MHD turbulence in the accretion disc or accretion disc rim.

The bipolar jet features observed in RXJ0513 and MR Vel are reminiscent of the Doppler shifted bipolar jet features in the Galactic microquasar SS 433, although the velocities in this source are about an order of magnitude larger than in the SSSs (e.g. Blundell et al., 2007). The presence of collimated jets in RXJ0513 and MR Vel, and also the less collimated outflow in CAL 83, indicates the possibility of observing radio emission from these jet-like structures. These “radio jets” have been observed in several Galactic and extragalactic systems, including SS 433 (e.g. Band and Grindlay, 1986; Hjellming and Johnston, 1988).

Although thermal radio emission may be observed from the cooler regions in these jets, non-thermal emission is often observed, which is also an important signature process related to magnetic fields in the system. When charged particles are ac-

celerated in a magnetic field, polarized radiation with a characteristic spectral energy distribution is produced (e.g. Rybicki and Lightman, 2004, Chapter 6). If the motion of the charged particles is ultra relativistic, the non-thermal emission is known as *synchrotron radiation*, which is often observed in radio jets. In jet regions, electrons can potentially be accelerated by shocks to relativistic velocities. In the non-relativistic case, the radiation is referred to as *cyclotron radiation*. Outflows provide ideal environments in which to observe non-thermal radiation, due to their low particle density and associated absorption.

CAL 83 and RXJ0513 have very low orbital inclinations, and the inclination of MR Vel is also relatively low. This provides the observer with a relatively unobstructed view of the outflow, more or less along the outflow axis. The LIMA-type model for the ~ 67 s periodicity in CAL 83 indicates the presence of a weakly magnetized WD in the system. The quasi-periodic optical modulations in all three sources, as well as the presence of outflows, can be readily explained by the presence of magnetic fields in the system. Thus, non-thermal emission from these sources is a distinct possibility.

Current radio data of these systems only provide an upper limit to the radio flux. However, once the SKA (Square Kilometre Array) is online, the unprecedented sensitivity of this instrument may be able to detect non-thermal emission from SSSs, enabling a better understanding of the outflows in these systems, as well as direct measurements of the magnetic field strengths for the first time. This may already be possible with the SKA precursor, the MeerKAT telescope array.

Once the polarimetry mode on SALT is available, polarization measurements of these sources will be undertaken, with the aim of constraining the magnetic field strength from the observed cyclotron radiation. However, with the optical emission being dominated by the accretion disc and envelope, it may be difficult to detect potential polarized cyclotron emission. If the magnetic field is weak, the cyclotron emission peak will be at longer wavelengths, and therefore be more visible in the infrared.

Another very interesting and extremely important branch of follow-up work on these SSSs will be to apply the simultaneous multiwavelength modelling techniques discussed by Skopal (2013, 2015a,b,c) to the existing data of the known supersoft X-ray source sample. This will enable more accurate constraints on the source parameters.

In this study, various fascinating properties related to variability in supersoft X-ray binaries have been uncovered. However, many questions have also been raised, and it is clear that the last word regarding these fascinating multiwavelength sources has

not been spoken, and they will remain a topic of considerable interest in the field of compact binaries.

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This research has made use of NASA’s Astrophysics Data System.

Appendix A

Physical constants

Table A.1: Fundamental physical constants. (Adopted from Zombeck, 1990, p. 10.)

Constant	Symbol	Value
Speed of light in vacuum	c	$2.998 \times 10^{10} \text{ cm s}^{-1}$
Gravitational constant	G	$6.672 \times 10^{-8} \text{ dyn cm}^2 \text{ g}^{-2}$
Planck's constant	h	$6.626 \times 10^{-27} \text{ erg s}$
Electron charge	e	$4.803 \times 10^{-10} \text{ esu}$
Electron mass	m_e	$9.109 \times 10^{-28} \text{ g}$
Proton mass	m_p	$1.673 \times 10^{-24} \text{ g}$
Boltzmann constant	k	$1.381 \times 10^{-16} \text{ erg K}^{-1}$
Stefan-Boltzmann constant	σ	$5.671 \times 10^{-5} \text{ erg cm}^{-2} \text{ K}^{-4} \text{ s}^{-1}$
Thomson cross-section	σ_T	$0.665 \times 10^{-24} \text{ cm}^{-2}$

Table A.2: Solar properties. (Adopted from Zombeck, 1990, p. 25.)

	Symbol	Value
Solar mass	$M_\odot$	$1.989 \times 10^{33} \text{ g}$
Solar radius	$R_\odot$	$6.960 \times 10^{10} \text{ cm}$
Solar luminosity	$L_\odot$	$3.826 \times 10^{33} \text{ erg s}^{-1}$

Appendix B

Optical Monitor magnitudes of CAL 83

Table B.1: Instrumental magnitudes of CAL 83 obtained with the Optical Monitor on board *XMM-Newton*.

Observation	Filter	BJD _{TDB} – 2454000 (mid-exposure)	Exposure time (s)	Instrumental magnitude
0500860301	U	288.50019	3000	16.3473 ± 0.0090
	U	288.53846	3000	16.3444 ± 0.0087
	U	288.58079	3700	16.2563 ± 0.0072
0500860401	U	334.15738	3500	15.5664 ± 0.0023
	U	334.20144	3500	15.7001 ± 0.0023
0500860501	U	379.59384	3000	16.0765 ± 0.0069
	U	379.63212	3000	16.1440 ± 0.0073
	U	379.67492	3780	16.0713 ± 0.0060
0500860601	U	429.41122	4400	16.1108 ± 0.0054
	U	429.46569	4400	16.0362 ± 0.0051
	U	429.52016	4400	16.0697 ± 0.0054
	U	429.57465	4400	16.0878 ± 0.0054
	U	429.62844	4280	16.0755 ± 0.0054
0500860701	U	482.08592	3000	15.4778 ± 0.0040
	U	482.12421	3000	15.4765 ± 0.0040
0506530201	B	545.53399	800	17.035 ± 0.018
	B	545.54680	800	17.042 ± 0.014
	UVW1	545.55961	800	15.1823 ± 0.0095
	UVW2	545.57242	800	14.772 ± 0.066
	UVW2	545.58569	880	14.867 ± 0.066
0506530301	B	560.29053	1200	16.6928 ± 0.0084
	B	560.30797	1200	16.5598 ± 0.0075
	B	560.32542	1200	16.6985 ± 0.0085
	UVW1	560.34285	1200	14.8603 ± 0.0056
	UVW1	560.36030	1200	14.8477 ± 0.0055
	UVW2	560.37772	1200	14.363 ± 0.043
	UVW2	560.39518	1200	14.383 ± 0.043
	UVW2	560.41262	1200	14.449 ± 0.045

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Table B.1 – continued from previous page

Observation	Filter	BJD _{TDB} – 2454000 (mid-exposure)	Exposure time (s)	Instrumental magnitude
0506530401	UVW2	560.43585	2200	14.414 ± 0.032
	B	567.76095	800	16.5813 ± 0.0094
	B	567.77376	800	16.5826 ± 0.0093
	UVW1	567.78657	800	14.7616 ± 0.0066
	UVW2	567.79938	800	14.305 ± 0.051
0506530501	UVW2	567.81266	880	14.330 ± 0.049
	B	573.03126	1900	17.142 ± 0.012
	UVW1	573.05738	2000	15.3102 ± 0.0068
	UVM2	573.08987	3000	14.892 ± 0.023
	UVW2	573.12814	3000	14.928 ± 0.037
0506530601	B	574.08486	1900	17.171 ± 0.010
	UVW1	574.13181	2000	15.3427 ± 0.0070
	UVM2	574.15851	2000	14.883 ± 0.028
	UVW2	574.18522	2000	14.853 ± 0.043
	B	575.79023	1000	17.040 ± 0.013
0506530801	UVW1	575.80534	1000	15.1011 ± 0.0080
	UVM2	575.82164	1200	14.677 ± 0.033
	UVW2	575.83907	1200	14.770 ± 0.053
	B	577.45456	1200	17.139 ± 0.013
	UVW1	577.48011	2600	15.2637 ± 0.0055
0506530901	UVM2	577.51373	2600	14.829 ± 0.024
	UVW2	577.55376	3700	14.779 ± 0.031
	U	577.61080	2800	15.8606 ± 0.0053
	U	577.64676	2800	15.8257 ± 0.0050
	U	577.68272	2800	15.8543 ± 0.0055
0506531001	U	577.71868	2800	15.8322 ± 0.0053
	B	578.29427	1200	17.090 ± 0.012
	UVW1	578.31692	2100	15.1779 ± 0.0058
	UVM2	578.34708	2500	14.792 ± 0.024
	B	579.98218	1000	16.998 ± 0.012
0506531201	UVW1	579.99732	1000	15.1078 ± 0.0080
	UVM2	580.01648	1700	14.689 ± 0.028
	UVW2	580.04261	2200	14.621 ± 0.036
	B	581.85346	1200	17.142 ± 0.013
	UVW1	581.87265	1500	15.2599 ± 0.0073
0506531301	UVM2	581.89760	2200	14.859 ± 0.027
	UVW2	581.93242	3200	14.762 ± 0.032
	B	691.12795	1000	17.558 ± 0.030
	UVW1	691.14308	1000	15.752 ± 0.032
	UVM2	691.16225	1700	15.344 ± 0.040
0506531501	UVW2	691.18549	1700	15.279 ± 0.060
	B	726.97739	1000	16.994 ± 0.017

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Table B.1 – continued from previous page

Observation	Filter	BJD _{TDB} – 2454000 (mid-exposure)	Exposure time (s)	Instrumental magnitude
0506531701	UVW1	726.99252	1000	15.0440 ± 0.0087
	UVM2	727.01170	1700	14.687 ± 0.028
	UVW2	727.03435	1600	14.618 ± 0.043
	UVW1	981.86299	2000	15.837 ± 0.023
	UVM2	981.90358	4400	15.463 ± 0.027
	UVM2	981.95805	4400	15.468 ± 0.027
	UVM2	982.01254	4400	15.478 ± 0.027
	UVM2	982.06702	4400	15.406 ± 0.026
	UVM2	982.14233	4400	15.480 ± 0.027
	UVM2	982.19680	4400	15.531 ± 0.028
	UVW2	982.25128	4400	15.442 ± 0.042
	UVW2	982.30576	4400	15.419 ± 0.042
	UVW2	982.34888	2440	15.454 ± 0.057

Table B.2: Specific flux measurements and AB magnitudes for CAL 83 obtained with the Optical Monitor on board *XMM-Newton*.

Observation ID	Filter	BJD _{TDB} – 2454000 (mid-exposure)	Specific flux f_λ (10^{-15} erg s $^{-1}$ cm $^{-2}$ Å $^{-1}$)	Magnitude m_{AB}
0500860301	U	288.53981	1.1626 ± 0.0051	17.245 ± 0.010
0500860401	U	334.17941	2.1829 ± 0.0033	16.561 ± 0.010
0500860501	U	379.63363	1.4217 ± 0.0051	17.027 ± 0.010
0500860601	U	429.52003	1.4496 ± 0.0032	17.005 ± 0.010
0500860701	U	482.10507	2.5158 ± 0.0065	16.407 ± 0.010
0506530201	B	545.54039	0.973 ± 0.010	16.853 ± 0.011
	UVW1	545.55961	3.102 ± 0.027	16.545 ± 0.010
	UVW2	545.57906	5.96 ± 0.26	16.524 ± 0.047
0506530301	B	560.30798	1.3931 ± 0.0060	16.463 ± 0.010
	UVW1	560.35157	4.197 ± 0.015	16.216 ± 0.010
	UVW2	560.40534	8.75 ± 0.17	16.107 ± 0.021
0506530401	B	567.76735	1.4810 ± 0.0090	16.397 ± 0.010
	UVW1	567.78657	4.570 ± 0.028	16.124 ± 0.010
	UVW2	567.80602	9.45 ± 0.31	16.023 ± 0.035
0506530501	B	573.03126	0.8846 ± 0.0098	16.956 ± 0.012
	UVW1	573.05738	2.757 ± 0.017	16.673 ± 0.010
	UVM2	573.08987	4.97 ± 0.11	16.531 ± 0.023
	UVW2	573.12814	5.39 ± 0.18	16.634 ± 0.037
0506530601	B	574.08486	0.8609 ± 0.0082	16.986 ± 0.010
	UVW1	574.13181	2.676 ± 0.017	16.705 ± 0.010
	UVM2	574.15851	5.01 ± 0.13	16.523 ± 0.028
	UVW2	574.18522	5.77 ± 0.23	16.558 ± 0.043
0506530801	B	575.79023	0.972 ± 0.011	16.855 ± 0.013
	UVW1	575.80534	3.343 ± 0.025	16.464 ± 0.010

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Table B.2 – continued from previous page

Observation ID	Filter	BJD _{TDB} – 2454000 (mid-exposure)	Specific flux f_λ (10^{-15} erg s $^{-1}$ cm $^{-2}$ Å $^{-1}$)	Magnitude m_{AB}
0506530901	UVM2	575.82164	6.06 ± 0.18	16.317 ± 0.033
	UVW2	575.83907	6.23 ± 0.31	16.475 ± 0.053
	B	577.45456	0.887 ± 0.011	16.954 ± 0.013
	UVW1	577.48011	2.878 ± 0.015	16.626 ± 0.010
	UVM2	577.51373	5.27 ± 0.12	16.469 ± 0.024
0500860901	UVW2	577.55376	6.18 ± 0.17	16.484 ± 0.031
	U	577.66474	1.7959 ± 0.0043	16.773 ± 0.010
0506531001	B	578.29427	0.927 ± 0.010	16.905 ± 0.012
	UVW1	578.31692	3.115 ± 0.017	16.540 ± 0.010
	UVM2	578.34708	5.45 ± 0.12	16.431 ± 0.024
0506531201	B	579.98218	1.009 ± 0.012	16.813 ± 0.012
	UVW1	579.99732	3.322 ± 0.024	16.470 ± 0.010
	UVM2	580.01648	6.00 ± 0.15	16.328 ± 0.028
0506531301	UVW2	580.04261	7.15 ± 0.24	16.326 ± 0.036
	B	581.85346	0.884 ± 0.010	16.957 ± 0.013
	UVW1	581.87265	2.888 ± 0.019	16.622 ± 0.010
	UVM2	581.89760	5.12 ± 0.13	16.499 ± 0.027
0506531501	UVW2	581.93242	6.28 ± 0.19	16.467 ± 0.032
	B	691.12795	0.603 ± 0.016	17.372 ± 0.030
	UVW1	691.14308	1.835 ± 0.055	17.115 ± 0.032
	UVM2	691.16225	3.28 ± 0.12	16.984 ± 0.040
0506531601	UVW2	691.18549	3.90 ± 0.22	16.984 ± 0.060
	B	726.97739	1.013 ± 0.016	16.809 ± 0.017
	UVW1	726.99252	3.524 ± 0.028	16.406 ± 0.010
	UVM2	727.01170	6.01 ± 0.16	16.326 ± 0.028
0506531701	UVW2	727.03435	7.17 ± 0.28	16.323 ± 0.043
	UVW1	981.86299	1.698 ± 0.037	17.199 ± 0.023
	UVM2	982.04672	2.918 ± 0.030	17.110 ± 0.011
	UVW2	982.30198	3.367 ± 0.085	17.143 ± 0.027

Appendix C

Properties of spectral lines

For a detailed discussion of the formation, broadening and properties of spectral lines, the reader is referred to the M.Sc. dissertation of the author (Odendaal, 2012, Chapter 3). In this Appendix, a few basic properties of spectral lines that are important for the discussion in Chapter 5 will be highlighted.

C.1 The line position

The observed wavelength position of the centre of a spectral line will be referred to as λ_{obs} , while the laboratory (rest) wavelength of the line is indicated by λ_0 . The relation between these two wavelengths yields information about the radial component of the velocity of the source relative to the observer, i.e. v_{obs} , according to the Doppler effect (also see §2.2):

$$v_{\text{obs}} = \frac{\lambda_{\text{obs}} - \lambda_0}{\lambda_0} c . \quad (\text{C.1})$$

The value of v_{obs} thus describes the motion of the source relative to the observer, but it is customary to rather use the solar system barycentre (SSB) as a much more stable reference point. The radial velocity of the emitting region relative to the SSB at the time of observation is

$$v_{\text{SSB}} = v_{\text{obs}} + \text{BCV} , \quad (\text{C.2})$$

where BCV represents the barycentric correction to the velocity, i.e. the speed of the observer towards the solar system barycentre (SSB) projected on the line of sight to the target. A positive BCV value indicates motion towards the target, and a negative value motion away from the target.

C.2 Ionization states

Ionization refers to the process during which a photon is absorbed by an electron in an atom, yielding an ion and a free electron. Roman numerals are usually used to indicate the degree of ionization. The number I refers to the neutral atom, and each succeeding number to one higher ionization state. E.g. O I refers to neutral oxygen, O III to O^{2+} , and O VI to O^{5+} .

The amount of energy that is required to ionize an atom from one state to another is the *ionization potential* ϵ_i . For an incident photon to cause ionization of an atom or ion, its energy must be equal to or greater than ϵ_i . The ionizing photons often originate in the local thermal radiation field, and in such a case the minimum required temperature can be derived from the ionization potential according to

$$T(\epsilon_i) = \epsilon_i \times \frac{1.602 \times 10^{-12} \text{ erg}}{1 \text{ eV}} \times \frac{1}{k} \approx 11600 \text{ K eV}^{-1} \times \epsilon_i, \quad (\text{C.3})$$

if ϵ_i is given in eV. The ionization potential of atoms and ions can be obtained e.g. on the NIST Atomic Spectra Database¹ (Kramida et al., 2014). Assuming that the radiation field is in equilibrium with the thermal motion of the particles in the gas, one can obtain a rough estimate for the temperature of the emitting region, i.e.

$$\frac{hc}{\lambda} \approx \frac{3}{2}kT. \quad (\text{C.4})$$

The equation above is valid for non-relativistic particles, and is therefore applicable to the emission regions in SSSs. Substituting the constants and rearranging yields

$$T(\lambda) \sim \frac{9.590 \times 10^7 \text{ K } \text{\AA}}{\lambda}. \quad (\text{C.5})$$

For $H\alpha$ photons, with a wavelength of 6563 $\text{\AA}$, the corresponding temperature according to the equation above is $\sim 14\,600$ K.

C.3 Line width and intensity

A spectral line has an approximately V-shaped structure, which is commonly known as the line “profile”. The width of the line at a height halfway between its base and maximum height is called its full width at half maximum (FWHM). Sometimes the half width at half maximum (HWHM) is also used, and is evidently given by

$$\text{HWHM} = \frac{1}{2} \text{FWHM}. \quad (\text{C.6})$$

¹<http://physics.nist.gov/PhysRefData/ASD/ionEnergy.html>

Another useful measure of the line width is its full width at zero intensity (FWZI), which is the width of the line measured at the continuum level, i.e. from the point λ_b on the blue side of the line centre where the line blends into the continuum, to the similar point λ_r on the red side.

The *equivalent width* of an emission line is a convenient measure of its intensity. Because the equivalent width expresses the line intensity relative to the continuum level, it can be determined even when an absolute flux calibration has not been applied to the spectrum. The equivalent width of a spectral line is given by (e.g. Allen, 1964, p. 164)

$$W_\lambda = \int_{\lambda_b}^{\lambda_r} \left(1 - \frac{I_\lambda}{C_\lambda}\right) d\lambda, \quad (\text{C.7})$$

where I_λ is the line intensity and C_λ the continuum intensity at wavelength λ . It is clear from this definition that an absorption line will have a positive equivalent width, while an emission line will have a negative equivalent width. In a practical sense, the equivalent width represents the required width that a rectangle with a height extending from zero to the average continuum level at the emission line should have, so that the area contained in the rectangle is equal to the area contained between the spectral line and the continuum level.

An appropriate equation for determining the error in an equivalent width measurement is provided by Vollmann and Eversberg (2006):

$$\sigma(W_\lambda) = \frac{(\text{FWZI} - W_\lambda)}{S/N} \sqrt{1 + \frac{\overline{F}_c}{\overline{F}}}. \quad (\text{C.8})$$

S/N represents the signal-to-noise ratio (S/N) of the spectrum, $\overline{F}_c$ the mean continuum flux and $\overline{F}$ the mean line flux.

C.4 Line broadening mechanisms

Because of the finite resolution of any spectrograph, the instrument itself will introduce a certain amount of broadening into an observed spectral line, known as *instrumental broadening*. A simple approximation that can be used to correct the observed full width at half maximum (FWHM_{obs}) for instrumental broadening is

$$\text{FWHM}_{\text{corr}} = \sqrt{\text{FWHM}_{\text{obs}}^2 - \text{FWHM}_{\text{inst}}^2}. \quad (\text{C.9})$$

$\text{FWHM}_{\text{inst}}$ represents the FWHM of the observed line profile that the instrument would record if the actual line emitted by the source was infinitely thin. $\text{FWHM}_{\text{inst}}$ can be taken to be approximately equal to the spectral resolution of the spectrograph.

According to the Heisenberg uncertainty principle,

$$\Delta E \Delta t \geq \frac{\hbar}{2}, \quad (\text{C.10})$$

i.e. the product of the uncertainty in an energy level and the lifetime of that energy level has a certain minimum value (e.g. Peach, 1975). Therefore, for a certain transition between energy levels in an atom, the associated photons will have a small spread in wavelengths, causing the observed spectral line to be slightly broadened. This is known as *natural line broadening*. In the classical approximation, the FWHM of any spectral line due to natural line broadening is 0.000118 Å. The natural line widths obtained with the more accurate quantum mechanical consideration are also usually completely overpowered by other line broadening mechanism, and are negligible in observational astronomical spectroscopy.

Because atoms contain positively charged protons and negatively charged electrons, any atom has an associated electric field. Atoms or ions undergoing a close encounter or “collision” will perturb each other’s electric fields, effectively perturbing the energies associated with certain energy levels and transitions within the atom or ion. As a result, a spectral line formed by transitions in a population of such perturbed atoms will have a broadened profile, an effect known as *pressure broadening* (e.g. Robinson, 2007, p. 61).

In reality, pressure broadening is a very complex phenomenon with several components, i.e. linear and quadratic Stark broadening, resonance broadening and Van der Waals interaction (see Lang, 1999, p. 188-196 for more information). Ion broadening by the linear Stark effect is the most important effect for hydrogen lines in the optical waveband, but electron broadening will also play a role, especially in cases where the transitions involve levels with large principal quantum numbers. Vidal, Cooper and Smith (1973) provided tables of Stark broadening values for hydrogen. A discussion of these tables was provided in Section 3.4.2 and Appendix C of Odenaal (2012).

Atoms or ions at any temperature T above absolute zero are in constant motion. The radial components of the random thermal motion of the emitting or absorbing atoms or ions will cause a distribution of Doppler shifts in the emitted or absorbed wavelengths. This also causes spectral line broadening, and the effect is called *thermal broadening*. A thermally broadened line has a Gaussian shape (e.g. Peach, 1975; Kitchen, 1995, Section 13.7.1), with a full width at half maximum of

$$\text{FWHM}_T = \sqrt{\frac{8kT \ln 2}{mc^2}} \lambda_0, \quad (\text{C.11})$$

where m is the mass of the atom or ion.

Other types of radial motion in astronomical sources also cause Doppler broadening, e.g. turbulence, rotation, expansion, contraction, winds and outflows. Such broadened profiles can sometimes have very interesting shapes, depending on the source geometry. For example, emission lines associated with accretion discs typically have a double-peaked structure (Smak, 1969; Horne and Marsh, 1986). In some systems, mass is ejected in collimated outflows or “jets”, and the spectral lines originating in the jet will be Doppler shifted from the main component of the same spectral line, as seen in e.g. some active galactic nuclei, as well as some SSSs (see Chapter 3).

Appendix D

CCD properties and reductions

Charge-coupled devices (CCDs) are nowadays the most common photon detectors in optical observational astronomy. Unfortunately, an image recorded with a CCD camera will not only contain signal from the target star, but also a variety of unwanted signals, e.g. the readout noise of the CCD, thermal current or hot pixels in the CCD, the effects of dust or internal reflections in the telescope, and also cosmic rays. Before performing any sensible astronomical measurements on the recorded data, these non-target related signal first needs to be removed, and the steps in this process are usually referred to as “CCD reductions”.

In this section, a selection of the basic properties of CCD cameras that are important to the observational astronomer are reviewed, especially as related to the necessary CCD reductions. This discussion is based on that of Howell (2006, pp. 1-87). A similar discussion can be found in Wodaski (2002, Chapter 6).

D.1 Quantum efficiency and band-pass

The detection of light by a CCD takes place by means of the photoelectric effect. An incident photon with appropriate energy is absorbed by the semiconductor material in the CCD, upon which a valence electron in the semiconductor is released and moves into the conduction band. In each pixel, the electrons freed in such a manner are collected in a potential well by a so-called gate structure until the end of the exposure, when the charges are read out.

The band-pass refers to the total spectral range over which the CCD is sensitive to photons, and this will be determined by, among others, the band gap energy of the semiconductor. For example, silicon has a useful photoelectric effect range of ~ 1.1 -10 eV. A photon with energy in the 1.1-4 eV range will produce a single electron-hole pair, while a photon with higher energy will produce multiple pairs.

The quantum efficiency of a CCD is defined as the percentage of the photons incident on the CCD that is actually detected. The quantum efficiency is usually given as a function of wavelength, and many modern research class CCDs have quantum efficiencies close to 100% at their sensitivity peak. Factors influencing the quantum efficiency include the composition and thickness of the CCD, as well as its operating temperature. Although a higher operating temperature (e.g. room temperature) is usually associated with a higher quantum efficiency, the dark current at high temperatures has a negative effect on the data quality, therefore CCDs are usually cooled to temperatures far below 0 °C (see §D.6).

D.2 Gain and readout noise

Every CCD system contains an analog-to-digital converter (ADC), the purpose of which is to convert the potential difference read out from each pixel to a digital unit (an analog-to-digital unit or ADU). The amount of electrons that will yield a digital signal of 1 ADU is called the “gain” of the CCD, for example 5 e⁻/ADU. The result is a two-dimensional CCD image in computer memory, where each pixel has a certain number of ADUs (also known as “counts”) associated with it.

However, the readout process itself imposes a certain amount of noise on the signal. Firstly, the conversion in the ADC is not perfectly repeatable, and multiple readouts of the same pixel voltage would produce a statistical spread of ADUs centred on a mean value. Secondly, the system electronics will also introduce spurious electrons, yielding unwanted variation in the output. These factors cause an additive uncertainty in the final ADU value for each CCD pixel, and are collectively known as “readout noise” or simply read noise.

D.3 Saturation, linearity and dynamic range

The largest ADU output value that a CCD can produce is determined by two factors:

- (i) The full well capacity of the CCD, which is the amount of charge that can be stored in a pixel at one time, and is typically of the order of 10⁵ electrons. Dividing the amount of electrons by the gain will yield the full well capacity in terms of ADUs.
- (ii) The number of bits in the ADC, e.g. a 16-bit ADC can produce output of up to 65 536 ADU. When one of these limits is reached, any additional incident photons will not produce additional ADUs, leading to inaccurate scientific measurements. This is known as “saturation”, and should always be avoided. Saturation may also cause photoelectrons to overflow or “bleed” to surrounding pixels.

The response of CCD cameras are linear over a large range, i.e. an increase in the amount of incident photons will cause a proportional increase in the ADU output

value. However, for some CCDs non-linearity sets in below the saturation limit. It is very important for the observer to be aware of the ADU value above which the CCD response is non-linear, so that this regime can be avoided.

(Even within the linear regime, other issues like charge diffusion and charge transfer inefficiency can also cause the ADU counts in some pixels to deviate slightly from the correct value, but these will not be discussed in detail here.)

A related concept is that of dynamic range, which describes the range over which the CCD is sensitive, defined as

$$D = \frac{\text{full well capacity}}{\text{read noise}} . \quad (\text{D.1})$$

In a practical sense, it represents the ratio between the brightest and faintest signals that can be measured.

D.4 Pixel size, binning and windowing

The size of a single pixel in a CCD camera is typically $\gtrsim 10 \mu\text{m}$. Of special interest to the observer is also the “plate scale” of the image, which is expressed in arcsec/pixel. For example, the CCDs used in the Robert Stobie Spectrograph have 4096 rows containing a ~ 8 arcmin spatial field of view, hence the plate scale of 0.13 arcsec/pixel.

CCD cameras allow what is called “on-chip” binning, during which the charge in a certain amount of pixels is added within the CCD output register, before read-out and analog-to-digital conversion. For example, if a certain CCD has a size of 1024×1024 pixels and the on-chip binning is set to 2×2 , the pixels will be regrouped into a grid of 512×512 “superpixels”, each consisting of a square containing 4 pixels. The charges accumulated within the 4 pixels in the superpixel will be combined before digitizing, therefore the original values of the separate charges will be lost.

Without binning, the value of a single pixel will be increased by $1 \times$ the readout noise during digitization. However, the readout and digitization of a binned superpixel will also add only $1 \times$ the readout noise, even though it contains $\sim 4 \times$ the signal of an individual pixel. Therefore binning can increase the signal-to-noise ratio (S/N), and will also decrease the readout time, although it evidently decreases the spatial resolution.

Many CCDs also provide the option of “windowing”, meaning that only a certain rectangular region (or regions) on the CCD is read out, allowing faster readout and also decreasing the amount of storage space needed for the data.

D.5 Overscan and bias

When a region of unexposed pixels on a CCD is read out, one would ideally expect all the pixels to have an associated ADU value of exactly zero. However, due to the nature of the readout noise, their values will have a small distribution about a mean of zero, including negative values. To avoid the occurrence of negative counts in the output, CCDs are set up to add a positive offset value to the whole image, known as the “bias level” or “zero level”, which may typically be ~ 400 ADU.

After obtaining science frames, the observer needs to remove this bias level from the image before proceeding with other calibration steps. This is usually performed in two steps, known as the overscan correction and bias correction respectively.

D.5.1 Overscan correction

Most CCD cameras store an extra number of rows or columns in the final image that have not been exposed to incident light. This is known as the overscan region, and does not represent physical pixels on the CCD, but rather pseudo-pixels generated just by sending additional clock cycles to the output electronics. Therefore they contain only the bias offset and readout noise, and can be used to remove the bias level.

Consider, for example, a certain 1024×1024 CCD image with an overscan region of 32 columns, i.e. a size of 32×1024 . There are three main ways of determining the appropriate overscan correction for each row. The simplest overscan subtraction would involve taking the mean or median of all 32 768 pixels in the overscan region, and subtracting this from all the other CCD pixels. However, this would not compensate for any two-dimensional non-uniformities that may have been introduced during the addition of the bias offset.

In the line-by-line method, the mean or median of the 32 pixels in each row in the overscan region is determined, and subtracted from all the exposed pixels in that row. The third method first averages the 32 overscan pixels in a row for all the rows respectively, providing a one-dimensional overscan vector consisting of 1024 values. A function (a constant or a polynomial with an appropriate order) is then fit to this vector, yielding a smoothed row-dependent overscan function. The value of this function at each row is evaluated, and then subtracted from all the exposed pixels along this row.

Both the line-by-line method and the smoothed overscan vector method will remove the bias level as well as any row-wise one-dimensional structure in the bias

level. However, to remove two-dimensional bias structure one needs to subtract bias frames, as described in the next section.

D.5.2 Bias correction

When a CCD image is obtained with an exposure time of 0 s, it represents only a readout of the whole unexposed CCD, and should contain only the bias level and readout noise. Such an image is known as a “bias frame” or “zero frame”, and will also contain the two-dimensional bias structure. When performing a bias correction, 10 or more bias frames should be obtained, and mean or median combined to create a “master bias”. After performing overscan subtraction for both the bias and target frames, a pixel-by-pixel subtraction of the master bias from the target frames will remove any residual two-dimensional structure that may have been present in the bias level.

D.6 Dark current

Thermal noise within the CCD will also cause electrons to be freed from the valence band and collected in the potential wells of the pixels involved, and once this occurs, these electrons can not be distinguished from those freed by photons from an astronomical object. This signal contamination is known as thermal current or “dark current”, as it will be present in any exposure, even if the shutter is closed and no light is incident on the CCD.

To minimize the dark current, science CCDs are therefore cooled by one of two methods. The first method involves the use of liquid nitrogen to cool the device to temperatures close to $-100\text{ }^{\circ}\text{C}$ (173 K). The second method is simpler and less expensive, and makes use of thermoelectric cooling methods to obtain CCD temperatures between $-20\text{ }^{\circ}\text{C}$ (253 K) and $-50\text{ }^{\circ}\text{C}$ (223 K).

D.6.1 Dark correction

The counts generated by dark current, as well as bad or “hot” pixels on the CCD, can be removed by using “dark frames”. These are CCD images taken with the shutter closed, blocking the incident light. Ideally, the exposure time for the dark frame should be equal to the exposure time used for your target frame. In such a case, the dark frame will contain not only the dark current and information about hot pixels, but also the two-dimensional bias level. Subtracting this dark frame from the target frame will simultaneously perform the bias and dark correction, and additional bias frames are not needed.

However, it is not always practical to obtain a dark frame with the same length as each and every one of the target frames, especially when observing faint targets requiring long exposure times. In such a case, it is assumed that the dark current increases linearly with exposure time, and the dark current is then scaled to the required exposure times during software processing. However, in this case the dark frame itself would first need to be overscan- and bias subtracted, otherwise the scaling would be erroneous.

One would in general obtain several dark frames, and after overscan and bias correction combine them to yield a master dark, which can be scaled to match the exposure time of the target image one wishes to correct. The target frame is then corrected by performing overscan and bias subtraction, and also subtracting the scaled master dark.

D.7 Flat fielding

In reality, each pixel on a CCD has a slightly different gain and quantum efficiency than the adjacent pixels. This has an adverse effect on the science data, as during analysis one would like to assume that all the pixels in the CCD have the same sensitivity to incident light. This effect is corrected by dividing the target frames by a “flat field”, which is an image obtained of a uniformly illuminated field, for example the twilight sky or an illuminated, reflective dome screen. The flat field division will not only remove the pixel-to-pixel variations in the sensitivity, but will also remove the effects of vignetting, internal reflections within the telescope, and dust motes that may be present anywhere in the optics.

Flat fielding is in principle a simple concept, but getting flat fielding really right in practice can be a very tough nut to crack. Firstly, the imaged field should be spatially flat, and the uniform illumination of every CCD pixel to less than 1% is very difficult to achieve.

Secondly, the imaged field should be spectrally flat, i.e. it should have the exact same wavelength distribution over the band-pass of interest as the target frames that need to be calibrated. This is because the quantum efficiency variations between the pixels are wavelength dependent. The twilight sky, or the quartz lamps used for dome flats, have spectral shapes very different from that of the dark star-filled sky. Sky flats of the dark sky are very time-consuming and absorb time that could have been used to observe astronomical targets, and are therefore rarely done.

However, twilight or sky flats provide acceptable flat fielding for many observational programmes. Ten or more flat frames are usually obtained, and bias and dark cor-

rected together with the target frames. Subsequently, the flat frames are combined to create a master flat, after which the target frames are divided by the master flat. The flat frames applied to a certain target image must be taken with exactly the same instrumental configuration as the target image, e.g. the same filter and, in the case of spectroscopy, the same grating.

D.8 Fringing

When performing observations in the red part of the optical spectrum, irregular dark and bright bands are often observed on a CCD image. They are essentially Newton’s rings, caused by thin film interference in the detector, and the effect is known as “fringing”. Light waves with longer wavelengths penetrate deeper into the silicon, and for wavelengths longer than a certain limit, the light waves can travel through the silicon array, and then be reflected back into the array, where it will undergo interference with other arriving light waves. For monochromatic light with an appropriately long wavelength, the result is a pattern of alternating brighter and darker bands due to constructive and destructive interference, with irregular ripple patterns due to small variations in the thickness of the silicon layer¹.

If the detector is illuminated by a broad-band light source in the red, the large range of wavelengths undergoing this effect will smear out the ripples, decreasing the fringing amplitude. An astronomical source with a broad-band spectrum will have a fringing pattern very similar to that of the broad-band illuminated flat field, therefore flat field subtraction can eliminate this fringing to a large extent. However, the sky is dominated by emission lines in the red part of the spectrum, primarily forbidden OH emission in the atmosphere, driven by UV sunlight absorbed during the day. These lines are very narrow, causing sharp fringes on the CCD image, and can also be variable during one night, making it very challenging to remove them, and other measures in addition to flat fielding need to be taken (see e.g. Broadfoot and Kendall, 1968; Wagner, 1992).

D.9 Bad pixels

In addition to the effects explained in the sections above, bad pixels may be present in a CCD image. A bad pixel is any pixel in the CCD that does not react as it should to incident light. Some pixels (or sometimes even an entire column) may be “dead”, i.e. have no response, while other pixels may be “hot”, i.e. always presenting an erroneous, very high count value. These can be corrected by identifying the bad pixels and then interpolating across them. Some observatories provide bad pixel

¹Also see <http://www.cfht.hawaii.edu/Instruments/Elixir/fringes.html>.

maps indicating the positions of the bad pixels on the relevant CCD, but one can also locate the bad pixels by a statistical analysis of flat fields.

Appendix E

Differential photometry method

In this appendix, a differential photometry method largely based on the method described by Everett and Howell (2001) will be discussed. This technique was applied to obtain corrected instrumental magnitudes CAL 83, RXJ0513 and MR Vel from the SHOC lightcurves in §6.1.

Let N be the number of comparison stars in the ensemble. The number of master lightcurves to be considered is $N + 1$: one for the target and one for each comparison. The parameter i will indicate the i^{th} lightcurve, with $i = 0$ representing the target, and $i = 1, 2, \dots, N$ representing the respective comparison lightcurves. The lightcurves all have the same length M , and the parameter $j = 1, 2, \dots, M$ indicates the frame number. It is emphasized that, for a particular value of j , the data points in the $N + 1$ lightcurves were all obtained from the same data frame. The data points in each lightcurve take the form $(t_{ij}, m_{ij}, \sigma(m_{ij}))$, where t_{ij} is the observation time at mid-exposure, m_{ij} the observed instrumental magnitude and $\sigma(m_{ij})$ its error.

Firstly, the mean magnitude error was calculated for each of the N comparisons:

$$\langle \sigma(m_i) \rangle = \frac{1}{M} \sum_{j=1}^M \sigma(m_{ij}) , \quad (\text{E.1})$$

for $i = 1, 2, \dots, N$. A mean comparison magnitude was consequently calculated for each frame j by averaging the comparison magnitudes for that particular frame, applying the inverse of its mean magnitude error $\langle \sigma(m_i) \rangle$ as a weight for each comparison:

$$\langle m_j \rangle = \frac{1}{S} \sum_{i=1}^N \frac{m_{ij}}{\langle \sigma(m_i) \rangle} , \quad (\text{E.2})$$

where

$$S = \sum_{i=1}^N \frac{1}{\langle \sigma(m_i) \rangle} \quad (\text{E.3})$$

represents the sum of the applied weights. The associated error in $\langle m_j \rangle$ is

$$\sigma(\langle m_j \rangle) = \sqrt{\sum_{i=1}^N \left(\frac{\sigma(m_{ij})}{S \times \langle \sigma(m_i) \rangle} \right)^2}. \quad (\text{E.4})$$

The overall mean of the comparisons over all the frames was also calculated:

$$\langle m_{\text{comp}} \rangle = \frac{1}{M} \sum_{j=1}^M \langle m_j \rangle. \quad (\text{E.5})$$

The differential correction term for frame j is now given by $\langle m_j \rangle - \langle m_{\text{comp}} \rangle$, which represents the deviation of the mean comparison magnitude for frame j from the constant reference magnitude $\langle m_{\text{comp}} \rangle$. The corrected instrumental magnitudes of the target can now be obtained by subtracting this differential term from the observed instrumental magnitudes: As a way of checking the quality of the correction, the same correction was applied to all the comparison lightcurves as well. The corrected instrumental magnitude is given by

$$m_{*ij} = m_{ij} - (\langle m_j \rangle - \langle m_{\text{comp}} \rangle), \quad (\text{E.6})$$

and its error is obtained by propagating the errors associated with each of the terms:

$$\sigma(m_{*ij}) = \sqrt{[\sigma(m_{ij})]^2 + [\sigma(\langle m_j \rangle)]^2}. \quad (\text{E.7})$$

The values of m_{*ij} and $\sigma(m_{*ij})$ were thus calculated for the target ($i = 0$) and all the comparisons ($i = 1, 2, \dots, N$). (Because the reference magnitude $\langle m_{\text{comp}} \rangle$ is simply the same constant for all the frames, its error was not included in the error calculation above.)

In an idealized scenario, the corrected magnitudes of each of the comparisons should now be perfectly constant over time. If only a single comparison was used, this will indeed be the case, by definition of the process. However, in reality, when several comparisons were used, their lightcurves will still exhibit a small spread in magnitudes around their respective mean values, the latter being given by

$$\langle m_{*i} \rangle = \frac{1}{M} \sum_{j=1}^M m_{*ij} \quad (\text{E.8})$$

for each of the comparisons ($i = 1, 2, \dots, N$). In principle, because the comparisons are assumed to be intrinsically constant, the statistical uncertainties in the corrected target instrumental magnitudes can not be smaller than the standard deviation of the comparison magnitudes from their respective mean values. To test this principle,

the mean error in the corrected target lightcurve was first calculated as

$$\langle \sigma(m_{*0}) \rangle = \frac{1}{M} \sum_{j=1}^M \sigma(m_{*0j}) \quad (\text{E.9})$$

which is simply a scalar value. The standard deviation (SD) in the corrected instrumental magnitudes was calculated for each of the comparisons, i.e.

$$\text{SD}_{*i} = \sqrt{\frac{1}{M} \sum_{j=1}^M (m_{*ij} - \langle m_{*i} \rangle)^2} \quad (\text{E.10})$$

and consequently averaged over all the comparisons, using the same weights that were applied previously to calculate the differential correction:

$$\text{SD}_{\text{comp}} = \frac{1}{S} \sum_{i=1}^N \frac{\text{SD}_{*i}}{\langle \sigma(m_i) \rangle} . \quad (\text{E.11})$$

The value of the mean target uncertainty $\langle \sigma(m_{*0}) \rangle$ was then compared to the value of the averaged standard deviation in the comparison magnitudes SD_{comp} . If the original error calculations determined by IRAF was realistic, then the value of $\langle \sigma(m_{*0}) \rangle$ should be comparable to or larger than SD_{comp} . If it was found that $\langle \sigma(m_{*0}) \rangle < \text{SD}_{\text{comp}}$, a multiplicative factor of $\text{SD}_{\text{comp}} / \langle \sigma(m_{*0}) \rangle$ was applied to all the individual error values $\sigma(m_{*0j})$ in the target lightcurve, to scale the latter values up to a more realistic uncertainty estimate.

Appendix F

Publications

During the course of this Ph.D. research (2012-2014), the following first-author MNRAS paper and two conference proceedings papers have been published:

- Odendaal, A., Meintjes, P. J., Charles, P. A. and Rajoelimanana, A. F. (2014), ‘Optical and X-ray properties of CAL 83 - II. An X-ray pulsation at ~ 67 s’, *MNRAS* **437**, 2948–2956.
- Odendaal, A., Meintjes, P. J., Charles, P. A. and Rajoelimanana, A. F. (2014*b*), Periodic X-ray modulations in Supersoft X-ray Sources, *in* J. Janse van Rensburg, ed., ‘Proceedings of SAIP2012, the 57th annual conference of the South African Institute of Physics’, Department of Physics, University of Pretoria, p. 356.
- Odendaal, A., Meintjes, P. J., Charles, P. A. and Rajoelimanana, A. F. (2013), ‘X-ray periodicity in the supersoft X-ray sources CAL 83 and SMC 13 .’, *Mem. S.A.It.* **84**, 545.

The following first-author conference proceedings papers have been submitted, but not yet published:

- Odendaal, A. and Meintjes, P. J. (2015*a*), ‘CAL 83 — The Prototypical Close Binary Supersoft X-ray Source in the LMC: A Short Review’, *Conference Papers in Science* **accepted**.
- Odendaal, A. and Meintjes, P. J. (2015*b*), ‘Multiwavelength variability in CAL 83’, *Mem. S.A.It.* **submitted**.
- Odendaal, A., Meintjes, P. J., Charles, P. and Rajoelimanana, A. F. (2015), ‘Supersoft X-ray Source CAL 83: A Possible AE Aqr-like System’, *Acta Polytechnica* **in press**.

The 6 first-author papers listed above are included in this appendix.

The author also acted as co-author on the following papers on projects closely related to the theme of this thesis:

- Meintjes, P. J., Odendaal, A. and van Heerden, H. (2015), ‘AE Aquarii: A Short Review’, *Acta Polytechnica* **in press**.
- Rajoelimanana, A. F., Charles, P. A., Meintjes, P. J., Odendaal, A. and Udalski, A. (2013), ‘Optical and X-ray properties of CAL 83 - I. Quasi-periodic optical and supersoft variability’, *MNRAS* **432**, 2886–2894.
- Meintjes, P. J., Oruru, B. and Odendaal, A. (2012), ‘The multi-frequency properties of AE Aquarii and its evolution from a high mass transfer history.’, *Mem. S.A.It.* **83**, 643.

Additionally, the author contributed to the spectroscopic studies forming part of the following publications on other projects:

- Meintjes, P. J., Nkundabakura, P., van Soelen, B. and Odendaal, A. (2014), Modelling the Multifrequency SED of AGN Candidates among the Unidentified EGRET and Fermi Gamma-Ray Sources, *in* ‘The Tenth International Workshop on Multifrequency Behaviour of High Energy Cosmic Sources, Palermo, Italy, 2013. Acta Polytechnica CTU Proceedings 1(1)’, pp. 163–169.
- Van Soelen, B., Meintjes, P. J., Odendaal, A. and Townsend, L. J. (2012), ‘Anisotropic inverse Compton scattering of photons from the circumstellar disc in PSR B1259-63’, *MNRAS* **426**, 3135–3147.



Optical and X-ray properties of CAL 83 – II. An X-ray pulsation at ~ 67 s^{*}

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ABSTRACT

CAL 83 is the prototypical close binary supersoft X-ray source in the Large Magellanic Cloud, has a 1 d orbital period, and is believed to consist of a white dwarf (WD) primary accreting from an evolved donor. Based on published WD model atmosphere fits to X-ray data, the WD has a mass of $\sim 1.3 M_{\odot}$, just below the Chandrasekhar limit. From a systematic search through archival *XMM-Newton* data for periodic emission from CAL 83 down to the shortest possible period just above the WD break-up period, we report the discovery of an ~ 67 s supersoft X-ray modulation, which we interpret as the rotation period of a highly spun-up WD. Such a short period can be explained within the framework of a high mass accretion history, where accretion disc torques could have spun up the WD over time-scales comparable to the thermal time-scale. The presence of carbon, oxygen and nitrogen in published optical and ultraviolet spectra may suggest CNO cycling in the envelope of a secondary star that is oversized for its inferred mass, suggesting that the secondary star shed a significant fraction of its envelope during a high mass-transfer history, resulting in a highly spun-up WD. The reported 67 s period shows an approximately ± 3 s drift from the median value in single runs, which we interpret as a hydrogen burning gas envelope surrounding the WD, with a period not quite synchronized with the WD rotation period.

Key words: accretion, accretion discs – binaries: close – stars: individual: CAL 83 – stars: oscillations – white dwarfs – X-rays: stars.

1 INTRODUCTION

Supersoft X-ray sources (SSSs) constitute a class of highly luminous astronomical sources that are characterized by strong emission in the supersoft X-ray band, with $\gtrsim 90$ per cent of the unabsorbed X-ray flux emitted below ~ 0.5 keV. Typical bolometric luminosities are $\sim 10^{36} - 10^{38}$ erg s⁻¹, which are close to the Eddington limit for a solar-mass object, and effective temperatures fall in the region $\sim 15 - 80$ eV (see Kahabka & van den Heuvel 2006, for a recent review on SSSs).

SSSs were first observed with the *Einstein X-ray Observatory* (Long, Helfand & Grabelsky 1981; Seward & Mitchell 1981) and further discoveries by *ROSAT* established them as a distinct class of objects (Trümper et al. 1991). Currently, more than 100 SSSs are known, of which most are in the Large and Small Magellanic Clouds and M31, with a few in the Milky Way Galaxy and about 20 external

galaxies (Kahabka 2006). The low number of SSSs discovered in our Galaxy is ascribed to the high degree of interstellar absorption of supersoft X-rays in the Galactic plane.

Because of the similarities of their spectra to those of low-mass X-ray binaries, it was initially suggested that SSSs may be accreting black holes (Smale et al. 1988; Cowley et al. 1990) or neutron star systems (Greiner, Hasinger & Kahabka 1991; Kylafis & Xilouris 1993). However, it was shown by Van den Heuvel et al. (1992) that the very low effective temperatures and high luminosities of these sources can be satisfactorily explained by the nuclear burning of hydrogen on the surface of a white dwarf (WD) accreting from a binary companion (the close binary supersoft source, or CBSS model). The accretion rate required for steady nuclear burning is $\sim (1-4) \times 10^{-7} M_{\odot} \text{ yr}^{-1}$. Such a high accretion rate can be sustained if the donor mass is comparable to or greater than the WD mass (Paczynski 1971; Savonije 1983), as this will cause the Roche lobe of the donor to shrink and drive mass transfer on the thermal time-scale of the donor.

Many SSSs are believed to contain accretion discs. However, the WD luminosity due to nuclear burning is much larger than the accretion luminosity of the disc, and some of the X-ray energy emitted by the luminous WD irradiates the surface of the disc.

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Therefore, the main role of the disc in terms of radiation is the reprocessing of X-rays from the WD, and this greatly exceeds the accretion energy released in the disc (Popham & Di Stefano 1996). As a result, the accretion disc is highly luminous in the optical waveband, and much brighter than in classical cataclysmic variables (CVs).

CAL 83 is a CBSS in the Large Magellanic Cloud and was one of the first SSSs discovered by the *Einstein X-ray Observatory* (Long et al. 1981; Brown et al. 1994). Based on the variation in velocity of the He II $\lambda 4686$ emission line, Crampton et al. (1987, hereafter C87) found that the mass ratio $q = M_2/M_{\text{WD}}$ is smaller than 1 for any value of the primary mass. This contradicts the mass ratio of $q > 1$ predicted for CAL 83 by Van den Heuvel et al. (1992), and also violates the condition for sustained mass transfer above the nuclear burning limit required to drive emission in SSSs.

CAL 83 was observed by the *Chandra X-ray Observatory* on four occasions during 1999 and 2001. Three of them were during the X-ray off-state, and one during the X-ray on-state. Lanz et al. (2005) made use of detailed line-blanketed non-local thermodynamic equilibrium (NLTE) models during a combined analysis of *XMM-Newton* data and the *Chandra* observation, and their results indicate a massive WD primary ($M_{\text{WD}} \sim 1.3 \pm 0.3 M_{\odot}$). According to an analysis by Suleimanov & Ibragimov (2003), considering the parameters of CAL 83 obtained from blanketing approximations and LTE model atmospheres based on *ROSAT* data, the mass of the WD should be $1.2\text{--}1.3 M_{\odot}$ if the photospheric radius is larger than the degenerate WD radius by a factor of 2. Considering the optical light curve and emission-line widths, the inclination of CAL 83 has been estimated to be in the range $i = 20^{\circ}\text{--}30^{\circ}$ (see Cowley et al. 1998 and references therein).

However, CAL 83 is not a persistent X-ray source and several X-ray off-states have been observed. It also exhibits long-term optical variability which is anticorrelated with the long-term X-ray variability. A detailed review and new results on this aspect are provided by Rajoelimanana et al. (2013) in our companion paper on CAL 83 (Paper I). These authors also derived a refined orbital period of $1.047\,529(1)$ d for CAL 83 from OGLE-III photometry. Previously, Schmidtke et al. (2004, hereafter S04) derived the following orbital ephemeris from MACHO Project photometry:

$$T_0 = \text{JD } 245\,1500.953 \pm 0.004 + 1.047\,568E \pm 0.000\,003 \text{ d.}$$

Schmidtke & Cowley (2006) found a periodic signal of 38.4 min in the *Chandra*-LETG data of CAL 83, which they ascribed to possible non-radial pulsations in the accreting WD, similar to those observed in some novae. No WD spin period for CAL 83 has been reported in the literature so far.

The optical spectrum of CAL 83 is dominated by Balmer and He II emission lines, presumably from the accretion disc (Pakull, Ilovaisky & Chevalier 1985; C87; Smale et al. 1988; Charles, Southwell & Livio 1997; Cowley et al. 1998; Oliveira & Steiner 2006). The He II $\lambda 4686$ emission line has a broad, variable wing structure that has also been observed in H β by several of these authors. Oliveira & Steiner (2006) reported an emission feature near the expected position of [N II] $\lambda 6548$ that may well be a blueshifted feature of H α , and in the spectrum of Charles et al. (1997) the H γ emission line also has an extended wing. In each particular spectrum, these wing structures are either all towards the blue side of the main components of these emission lines, or all towards the red side. According to C87, this may possibly be associated with the slow precession (with a period ~ 69 d) of an accretion disc which exhibits outflow through either a wind or a jet with a large collimation angle. This possibility is supported by the arguments of Southwell,

Table 1. Archival *XMM-Newton* observations of CAL 83.

Observation ID	Start date and time (UT)		Exposure time (s)	Mean PN (counts s $^{-1}$)
0123510101	2000-04-23	07:34:01	45021	4.8
0500860201	2007-05-13	22:03:32	11949	6.4
0500860301	2007-07-06	23:31:52	10920	5.9
0500860401	2007-08-21	15:11:21	7915	4.6
0500860501	2007-10-05	23:49:21	17615	4.4
0500860601	2007-11-24	21:10:14	23173	5.7
0500860701	2008-01-16	13:24:33	10914	Off
0500860801	2008-03-10	10:14:38	6916	Off
0500860901	2008-03-21	02:10:24	12716	0.3
0506530201	2008-03-20	00:33:22	7715	0.8
0506530301	2008-04-03	18:40:55	14918	Off
0506530401	2008-04-11	06:02:24	5914	Off
0506530501	2008-04-16	12:32:01	10873	1.8
0506530601	2008-04-17	13:40:13	11217	1.5
0506530801	2008-04-19	06:43:35	5916	0.5
0506530901	2008-04-20	22:38:43	11617	0.5
0506531001	2008-04-21	18:47:57	9076	0.9
0506531201	2008-04-23	11:20:22	7418	0.3
0506531301	2008-04-25	08:13:32	9614	0.6
0506531401	2008-04-29	00:40:23	14117	0.3
0506531501	2008-08-12	14:50:27	6918	8.0
0506531601	2008-09-17	11:10:21	6817	0.2
0506531701	2009-05-30	08:00:48	46115	8.0

Livio & Pringle (1997), who showed that irradiation-induced accretion disc warping in CAL 83 can cause disc precession on these time-scales.

The paper will be structured as follows. In Section 2, results of X-ray (*XMM-Newton*) timing analysis will be presented. In Section 3, the possibility of the ~ 67 s X-ray pulse being associated with the spin period will be evaluated, and the conclusions are summarized in Section 4. The presence of an ~ 67 s oscillation was briefly reported at the conference ‘X-ray Astronomy: towards the next 50 years’ (Odendaal et al. 2013).

2 XMM-Newton X-RAY OBSERVATIONS AND RESULTS

CAL 83 has been observed by *XMM-Newton* several times from 2000 April to 2009 May. The X-ray data for all the observations were recorded with the three EPIC detectors (MOS1, MOS2 and PN) (Strüder et al. 2001; Turner et al. 2001) and the Reflection Grating Spectrometers (RGSs). All the observations, except 0123510101, also include data obtained with the Optical Monitor (OM). The observations are summarized in Table 1, where the mean count rate of the PN detector in the 0.15–2.5 keV band is also provided in order to compare the relative strength of the X-ray signal in the different observations.

All the data sets were calibrated by following standard data reduction procedures with the *XMM-Newton* SAS (Science Analysis System),¹ Version 12.0.1, and the arrival times in the calibrated event files were corrected to the Solar system barycentre (in Barycentric Dynamical Time or TDB). From each calibrated X-ray event file, a light curve with a binning of 5 s was created from the source photons in the 0.15–2.5 keV range. A source aperture with a radius of 40 arcsec was used. An OM light curve with a binning of 10 s was also generated for every observation with OM data.

¹ xmm.esa.int/sas

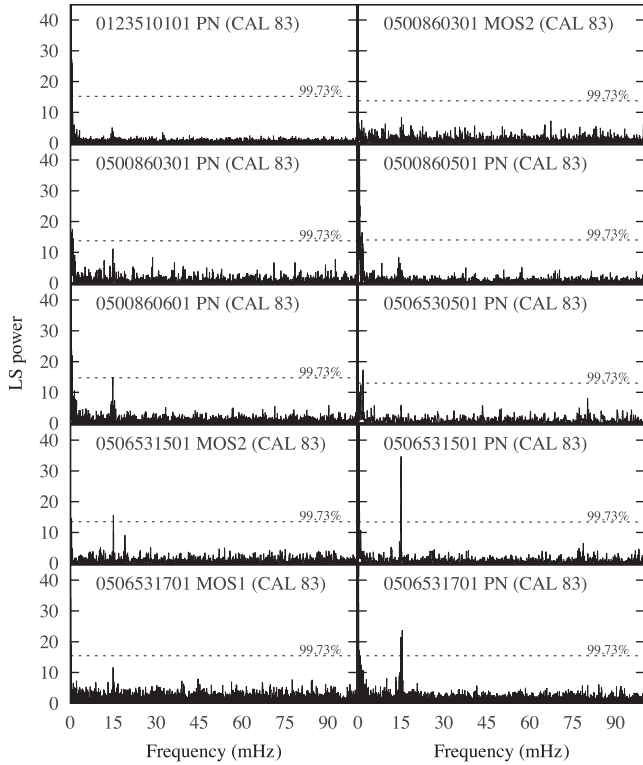


Figure 1. LS periodograms of CAL 83 *XMM-Newton* light curves with 15 mHz (~ 67 s) pulsation. The 99.73 per cent significance level, calculated over the whole 0–100 mHz range of each periodogram, is indicated.

A systematic search for short time-scale periodicities possibly associated with a WD spin was performed on all the *XMM-Newton* data sets. The timing analysis was performed with the Starlink PERIOD² package. Each light curve was first detrended by subtracting a linear fit and dividing by the standard deviation. The PERIOD task SCARGLE was then utilized to create a Lomb–Scargle (LS) periodogram from each light curve (Lomb 1976; Scargle 1982).

A peak near 15 mHz (~ 67 s) was discovered at a very high significance level (>99.9999 per cent) in the LS periodogram of the PN detector of observation 0506531501, and at a slightly lower level in six of the other PN data sets. In three of these seven observations, one of the MOS light curves also had a weak peak at this position. No significant peak was found at ~ 67 s in any of the RGS or OM data sets. The LS periodograms of the data sets exhibiting the periodicity are shown in Fig. 1. The statistical significance levels for Fig. 1 were determined by performing Monte Carlo significance tests over the whole range of frequencies from 0 to 100 mHz, including the red noise region. The LS analysis and associated Monte Carlo significance tests were subsequently performed for only the 10–20 mHz band, and the results are shown in Fig. 2. The results of this analysis are provided in Table 2, where the position of the strongest peak near 15 mHz is given, as well as its statistical significance as determined by the Monte Carlo significance tests over the 10–20 mHz band. There is no direct correlation between the significance of the peaks and the mean strength of the X-ray signal.

The corresponding folded light curves are shown in Fig. 3. Each respective light curve was folded on the period determined for it with PERIOD (given in Table 2), and a sine function with this period

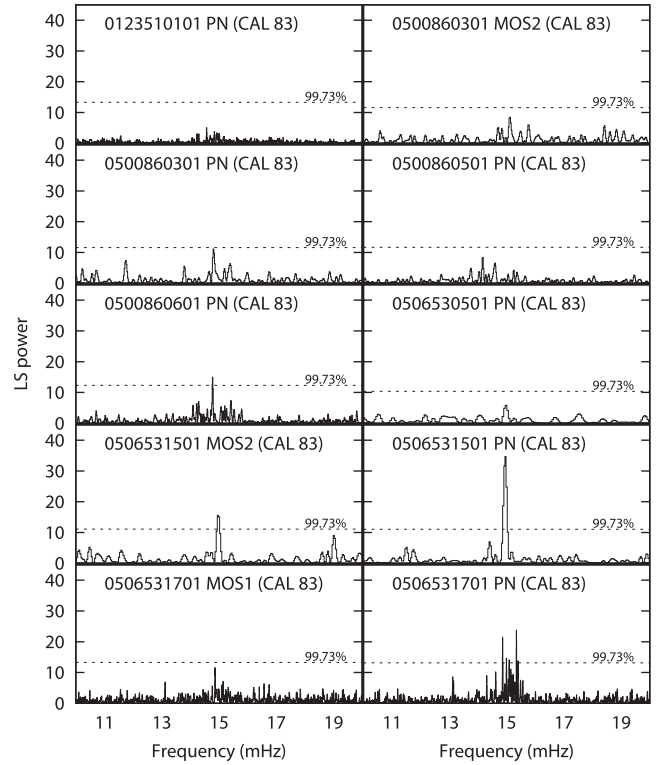


Figure 2. LS periodograms of CAL 83, showing the 10–20 mHz range and a closer view of the ~ 15 mHz peaks. The 99.73 per cent significance level is indicated, calculated over the 10–20 mHz interval. The values corresponding to the strongest ~ 15 mHz peaks are given in Table 2.

Table 2. Temporal analysis of CAL 83 *XMM-Newton* light curves showing the ~ 67 s pulsation.

Observation ID and instrument	Period (s)	Significance (%)	Modulation amplitude (%)
0123510101 PN ^a	68.64 ± 0.06	0.49	2
0500860301 MOS2	66.1 ± 0.2	94.50	6
0500860301 PN	67.5 ± 0.2	99.57	4
0500860501 PN	70.6 ± 0.2	92.37	4
0500860601 PN	67.7 ± 0.1	99.98	4
0506530501 PN	66.8 ± 0.5	75.10	8
0506531501 MOS2	66.8 ± 0.4	99.998	10
0506531501 PN	66.8 ± 0.4	>99.9999	7
0506531701 MOS1	67.28 ± 0.05	98.86	4
0506531701 PN	65.18 ± 0.05	>99.9999	2

^aEven though the weak power peak in the 0123510101 PN data set is not statistically significant according to the Monte Carlo significance tests, the modulation is visible in the folded light curve in Fig. 3.

fixed was fitted and overplotted. The light curves show a clear quasi-sinusoidal modulation. The modulation amplitude in the count rate for each folded light curve is also given in Table 2, and represents the ratio of the semi-amplitude of the sine fit to its average level.

A similar analysis was performed on the events in the background regions, and the corresponding LS periodograms are shown in Fig. 4. In addition to analysing the defined background regions, we also investigated possible contamination by counts from other sources by performing the analysis on all the detected off-source photons in the EPIC PN, MOS1 and MOS2 event files. However, no significant power peaks are present at ~ 67 s in any of these power spectra.

² www.starlink.rl.ac.uk/star/docs/sun167.htm#sun167.html

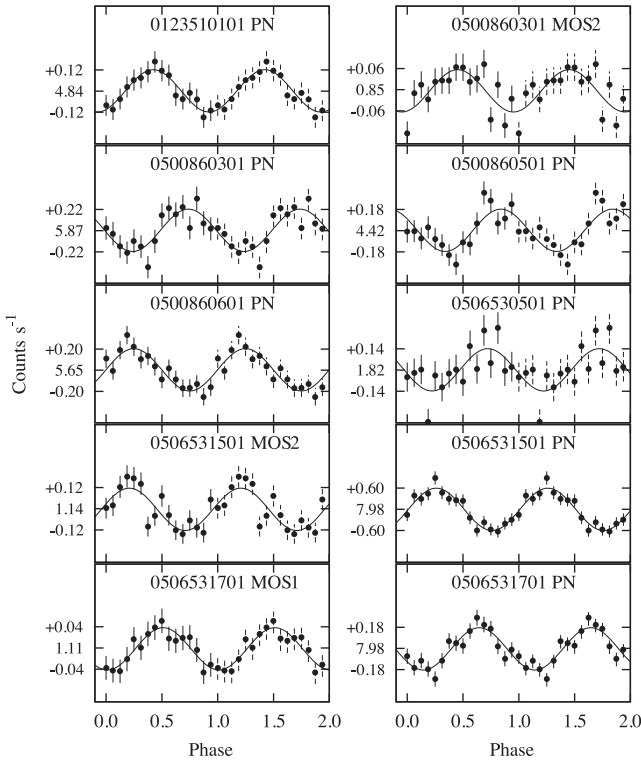


Figure 3. *XMM-Newton* light curves of CAL 83 folded on the corresponding periods in Table 2. The time of the first data point in each respective observation was used as the zero-point in each case. The vertical axis values represent the average level and semi-amplitude of the sine fit.

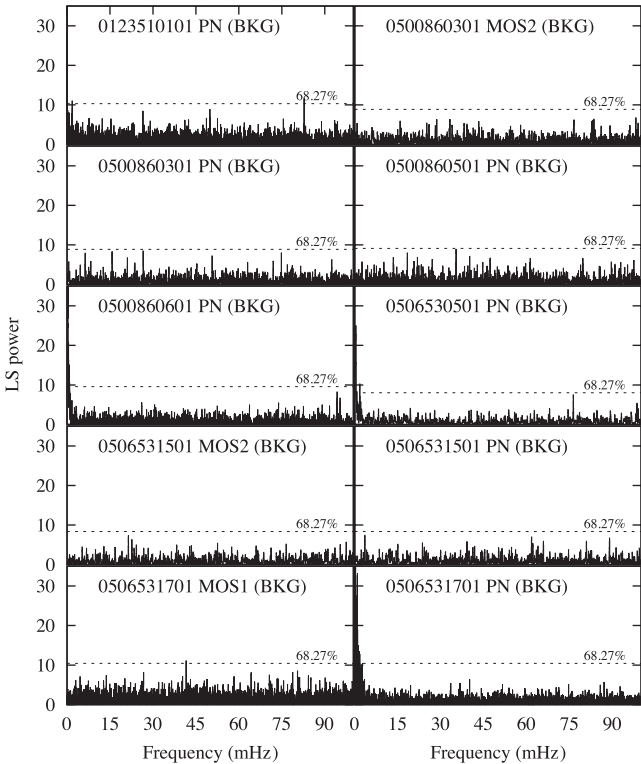


Figure 4. LS periodograms of background (BKG) light curves. The 68.27 per cent significance level is indicated, calculated over the 0–100 mHz range.

We especially considered contamination from the accreting pulsar XMMU J054134.7–682550, which is situated ~ 4 arcmin away from CAL 83. Although its position is outside the small sky region covered by the EPIC PN chip in Small Window mode, its position is included in the MOS1 and MOS2 detector windows. Manousakis et al. (2009) found pulsed emission from this source at a period of ~ 61 s in *XMM-Newton* observations 0500860301, 0500860401, 0500860501 and 0500860601, during which the source was observed in outburst. These authors also report that the detected spin period of XMMU J054134.7–682550 decreased by 1.54 s during ~ 50 d as a result of spin-up effects. By careful investigation we have excluded the possibility that the ~ 67 s oscillation arises from contamination by XMMU J054134.7–682550 photons, because (i) the power spectra of this pulsar calculated from the observations under consideration do not exhibit a peak corresponding to ~ 67 s, and (ii) the ~ 61 s period (which is decreasing due to spin-up effects) reported for the pulsar is easily resolvable from our ~ 67 s peaks.

Note that, for observations 0500860301, 0506531501 and 0506531701, respectively, the PN and MOS light curves were trimmed slightly at the edges so that they represent exactly the same time frame, and their binning intervals are also identical. Even so, the strongest power peaks in the PN and MOS periodograms of observations 0500860301 and 0506531701, respectively, are not included in each other's error ranges, as can be seen in Table 2. However, the 0506531701 PN periodogram in Fig. 2 shows that there are multiple peaks present at ~ 15 mHz, and that the second strongest one seems to be the strongest in the 0506531701 MOS1 periodogram. A similar argument can be made for the two 0500860301 periodograms. Therefore, even though the detectors are observing the source simultaneously, it seems that sampling differences may cause different peaks in our periodograms to be emphasized.

In order to investigate the overall significance of the periodicity at ~ 15 mHz, LS periodograms with a fixed frequency interval of 0.05 mHz were produced for both the source and background regions in all the on-state observations in Table 1. The average of these 19 periodograms for the source and background regions, respectively, was calculated and is shown in Fig. 5. The strong peak at 66.67 s is clearly visible above the noise level in the average periodogram for CAL 83, but is absent in the average background periodogram. To estimate its statistical significance, the standard deviation (σ) of the 66.67 s peak relative to the average white noise level in the 20–100 mHz band was determined, and was found to be at a $\sigma \geq 3$ level.

It is evident that there is significant variation in the position of the peak in the different periodograms of CAL 83 in Fig. 2, and also substantial variation in the peak power. In order to investigate this, as well as the presence of multiple peaks in most of the periodograms, the light curves with 5 s binning were divided into shorter segments, each with a length of approximately 2.5 ks, and the analysis was repeated for each segment. The LS periodograms for all the segments are shown in Fig. 6, and all the period detections above a 68.27 per cent significance level are indicated.

In Fig. 6, there are 27 PN segments with a period detection with >68.27 per cent significance, and these are listed in Table 3. The histogram in Fig. 7 provides a visual representation of the spread of these periods.

Clearly, the peak position and power are highly variable even within a single data set. For example, in the data set 0506531701 PN, which represents the longest observation extending over ~ 0.5 of an orbital cycle of CAL 83, the peak position in the ~ 2.5 ks

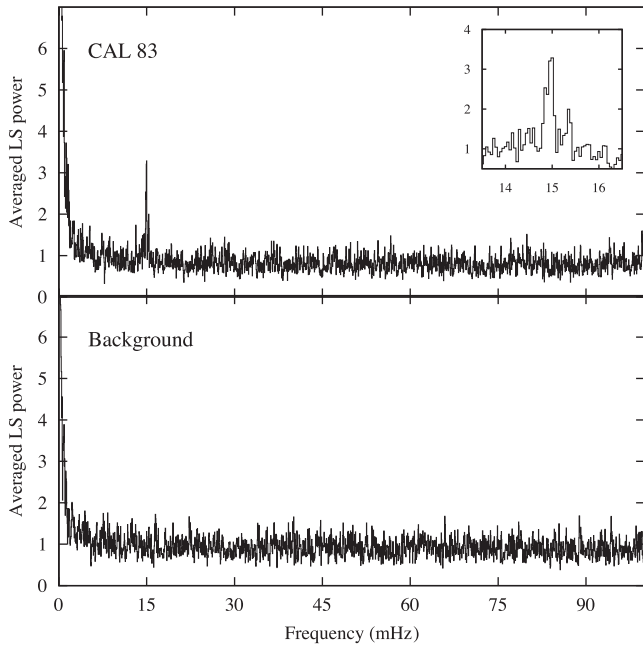


Figure 5. Average of LS periodograms of all on-state PN light curves (see Table 1) of CAL 83 and the background regions. The strong peak in the CAL 83 periodogram corresponds to a period of 66.67 s, and the inset shows its profile in more detail. The significance of this peak is estimated to be $\geq 3\sigma$.

segments varies between 64.4 ± 0.8 and 69.8 ± 1.0 s, and the significance from none up to a >99.999 per cent level.

The light curve of the 0506531701 PN data set consists of 9100×5 s bins. To investigate the variation in the detected period over the extent of this data set in more detail, this light curve was further divided into 2.25 ks light curves (i.e. 450×5 s bins) in such a way that each consecutive segment is shifted by only 100 s relative to the previous one, yielding a set of 433 light curves. An LS analysis was performed for each 2.25 ks light curve, the results of which therefore represent a ‘moving average’ of the value of the detected period throughout the 0506531701 PN data set, with an error of ± 1 s on each period determination. There were 281 period detections at a >68.27 per cent significance among the 433 light curves, and these are plotted in Fig. 8.

To explore the possibility of a relationship between the periodicity and the hardness of the light curve, the counts of each light curve were divided into two bands: the S band (0.15–0.3 keV) and the H band (0.3–2.5 keV). The hardness ratio (HR) was defined as $HR = H/(H + S)$, and an HR light curve was created for each of the 10 data sets. In Fig. 9, the S, H and HR light curves are shown for each respective data set, folded on the PN period of the observation given in Table 2. An LS analysis was also performed on the HR light curves, and the resulting periodograms are shown in Fig. 10. Upon inspection, some of the HR light curves in Fig. 9 seem to exhibit a modulation with the ~ 67 s period. However, no ~ 67 s periodicity can be claimed from the HR LS periodograms in Fig. 10.

Light curves with binning values down to 1.5 s have also been created for all the observations listed in Table 1. LS analyses similar to those used to obtain Figs 1 and 5 were also performed for these light curves, so as to investigate the possibility of periodicities at even higher frequencies (up to 333 mHz). However, no consistent periods except that at ~ 67 s were found.

3 A SHORT WD SPIN PERIOD?

The long-term presence of an ~ 67 s X-ray pulsation in CAL 83 suggests that it may be associated with the rotation period of the accreting WD. It has been observed in individual runs and verified at a significance level of $\geq 3\sigma$ by an incoherent combination (stacking) of the periodograms of all the on-state observations. This short period is reminiscent of the short spin periods in CVs like AE Aqr (33.08 s), V533 Her (63.633 s) and DQ Her (142 s) (e.g. Norton, Wynn & Somerscales 2004). The short rotation period in AE Aqr has been explained in terms of a high mass accretion history where the WD was spun up by accretion disc torques during intervals of high-mass-transfer rate, during which it might then appear as an SSS as a result of stable nuclear burning on the surface of the WD (Meintjes 2002; Schenker et al. 2002). The presence of supersoft X-ray emission in CAL 83 is also due to steady nuclear burning as a result of close to Eddington-limited accretion on to the surface of the WD (Van den Heuvel et al. 1992), placing CAL 83 in a similar situation to AE Aqr approximately ~ 10 Myr ago.

In the supersoft source population, the source XMMU J004252.5+411540 in M31 has exhibited a stable ~ 217.7 s supersoft X-ray modulation over a period of 5 yr that is ascribed to the spin period of the magnetic WD (Trudolyubov & Priedhorsky 2008). Similarly, the nova M31N 2007–12b exhibited a 1110 s stable X-ray pulsation during its supersoft phase, which is interpreted as the spin period of the WD in an intermediate polar system (Pietsch et al. 2011). Osborne et al. (2001) detected an ~ 865 s supersoft oscillation from the transient source XMMU J004319.4+411759, which is also interpreted as the spin period of an accreting WD. During the supersoft phase of its 2006 eruption, the recurrent nova RS Oph was reported to exhibit a multiperiodic X-ray modulation of ~ 35 s that became increasingly coherent (see Osborne et al. 2011, and references therein). These authors ascribed the modulation to non-radial pulsations driven by nuclear burning instabilities.

The ~ 67 s X-ray pulse in CAL 83 exhibits a drift in an approximately ± 3 s period band around 67 s on time-scales of hours. Due to the large moment of inertia of a WD, this drift cannot be the result of accretion-induced spin-up and spin-down (e.g. Frank, King & Raine 2002, p. 163). It can also be shown that the drift in the ~ 67 s pulse is not associated with the periodic effect introduced by orbital motion. The Doppler shift of the rotation period introduced by orbital motion can be constrained by the semi-amplitude (K_1) of the radial velocity curve associated with the position of a spectral line that originates on, or very close to, the accreting WD. Using a refined orbital ephemeris, S04 reanalysed the spectroscopic data of C87 and Smale et al. (1988), and determined the semi-amplitude of the He II $\lambda 4686$ emission line to be 27.0 ± 1.6 km s $^{-1}$. Utilizing this semi-amplitude and assuming a circular orbit, the associated drift of an ~ 67 s pulse as a result of orbital motion is

$$\Delta P = 0.0060 \left(\frac{K_1}{27.0 \text{ km s}^{-1}} \right) \left(\frac{P_*}{67 \text{ s}} \right) \text{ s.} \quad (1)$$

This is significantly smaller than the ± 3 s drift detected in individual runs. It is, however, important to keep in mind that measurements of K_1 from the radial velocity modulation of the He II $\lambda 4686$ accretion disc emission line assume that this line provides an exact measure of the WD orbital motion, which would be true if the He II emission was distributed perfectly symmetrically about the WD. Unfortunately, there are serious problems with this assumption, as discussed by, e.g. Cantrell & Bailyn (2007), compromising the accuracy of ΔP above, as well as the calculation of the mass function. However, even increasing the adopted value of K_1 by, e.g. an order

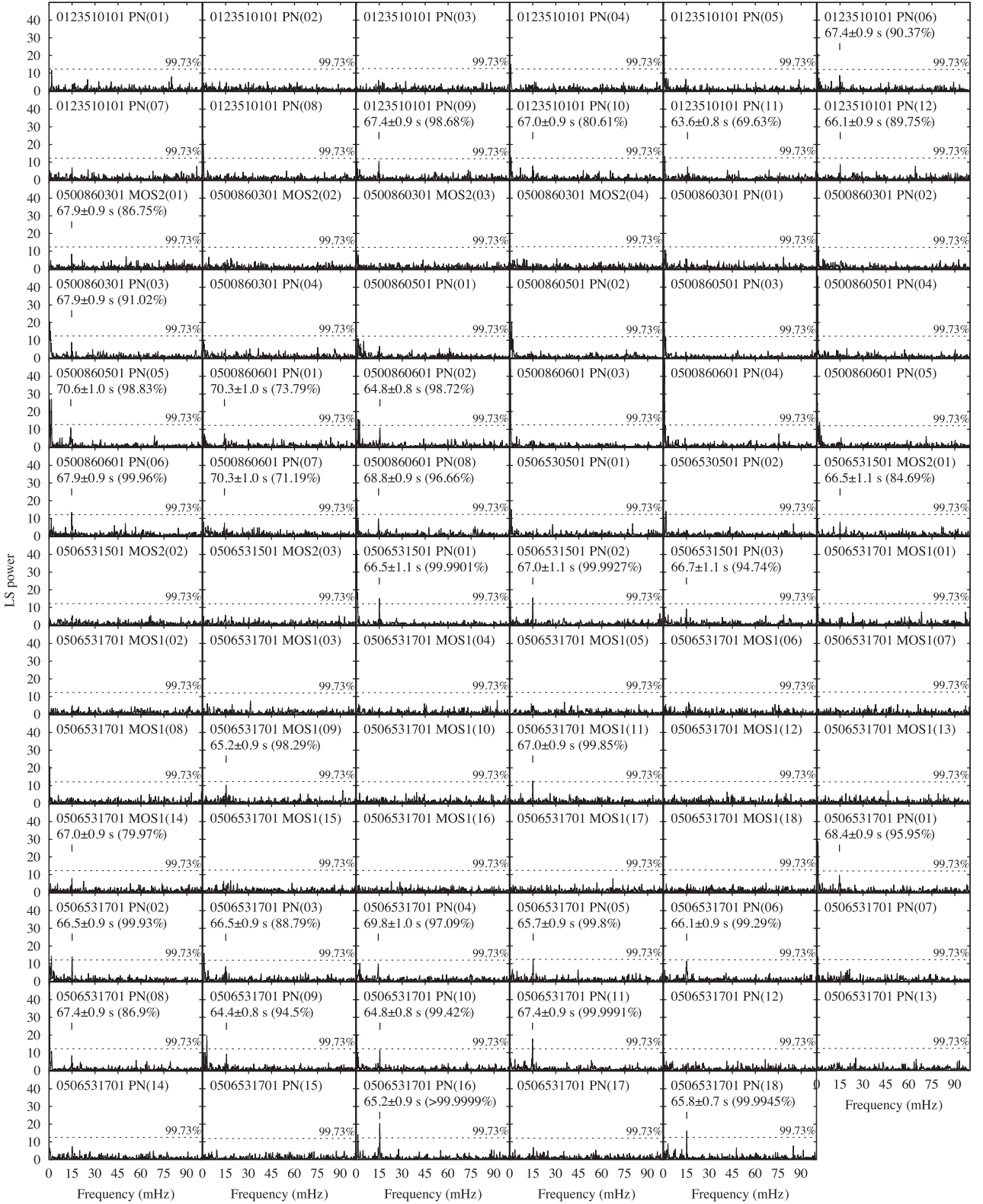


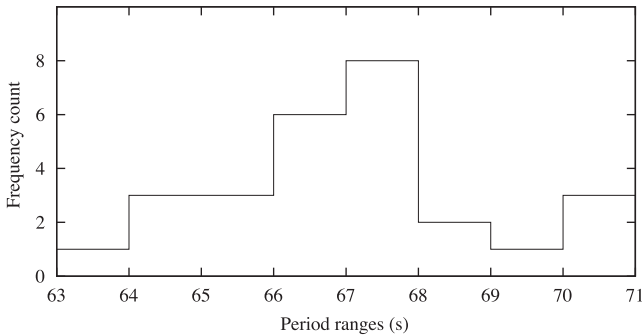
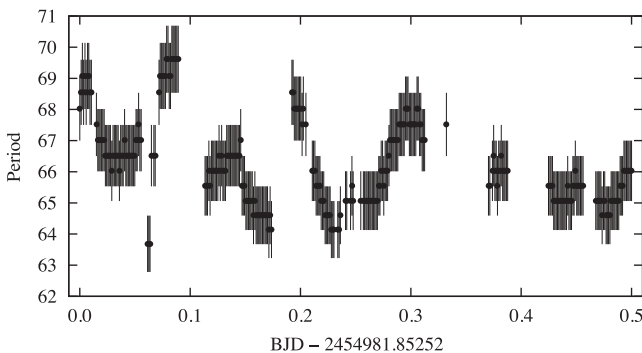
Figure 6. LS periodograms of those *XMM-Newton* EPIC light curves of CAL 83 that exhibit the 15 mHz (~ 67 s) pulsation, subdivided into ~ 2.5 ks segments. The peak position and significance are indicated only for those segments with a detection of >68.27 per cent.

of magnitude in equation (1) will not be able to explain the observed changes in the rotation period. Furthermore, these changes (see Fig. 8) are not smooth and do not correlate with the orbital phase.

Evidence exists indicating that CAL 83 has a long mass-transfer history, during which the WD could have been spun up to a short rotation period. Optical and ultraviolet (UV) spectra of CAL 83 contain emission lines of ionized carbon (C), nitrogen (N) and

Table 3. CAL 83 pulse period determinations.

Data set (segment)	BJD – 2450000 (mid-exposure)	Period (s)
0123510101 PN (06)	1658.08347	67.4 ± 0.9
(09)	1658.22698	67.4 ± 0.9
(10)	1658.25592	67.0 ± 0.9
(11)	1658.28485	63.6 ± 0.8
(12)	1658.31379	66.1 ± 0.9
0500860301 PN (03)	4288.55855	67.9 ± 0.9
0500860501 PN (05)	4379.68245	70.6 ± 1.0
0500860601 PN (01)	4429.43317	70.3 ± 1.0
(02)	4429.46210	64.8 ± 0.8
(06)	4429.57784	67.9 ± 0.9
(07)	4429.60678	70.3 ± 1.0
(08)	4429.63589	68.8 ± 0.9
0506531501 PN (01)	4691.13627	66.5 ± 1.0
(02)	4691.16058	67.0 ± 1.1
(03)	4691.18488	66.7 ± 1.1
0506531701 PN (01)	4981.85399	68.4 ± 0.9
(02)	4981.88293	66.5 ± 0.9
(03)	4981.91186	66.5 ± 0.9
(04)	4981.94080	69.8 ± 1.0
(05)	4981.96973	65.7 ± 0.9
(06)	4981.99867	66.1 ± 0.9
(08)	4982.05654	67.4 ± 0.9
(09)	4982.08547	64.4 ± 0.8
(10)	4982.11441	64.8 ± 0.8
(11)	4982.14334	67.4 ± 0.9
(16)	4982.28802	65.2 ± 0.9
(18)	4982.34878	65.8 ± 0.7

**Figure 7.** Histogram of the distribution of all the >68.27 per cent period detections according to the LS periodograms in Fig. 6 (only for PN data sets).**Figure 8.** Variation of the pulsation period with time within the duration of the 0506531701 PN data set. The horizontal axis represents the BJD of the middle of each 2.25 ks segment relative to the start of the observation.

oxygen (O) (see e.g. the spectra of C87, Gänsicke et al. 1998 and S04). C87 also commented that there is a low ratio of hydrogen to He II and CNO lines in the CAL 83 spectra, indicating a secondary with an H-poor envelope. This may indicate an evolved secondary that has shed most of its envelope in the mass transfer to the compact object, resulting in a highly spun-up WD. This evidence of CNO cycling of the secondary star envelope, as well as the WD mass being just below the Chandrasekhar limit, indicates that CAL 83 has a long mass accretion history.

For $M_{\text{WD}} \sim 1.3 M_{\odot}$ and using $P_{\text{orb}} = 1.047568$ d and $K_1 = 27.0 \text{ km s}^{-1}$ from S04, the secondary mass could be as low as $M_2 \sim 0.44 M_{\odot}$ (assuming that this K_1 represents the WD motion, which may be problematic as mentioned before). However, with a binary period of $P_{\text{orb}} \sim 1$ d, the Roche lobe radius of the secondary should be of the order of $R_2 \sim 1.5 R_{\odot}$. Therefore, if CAL 83 does possess a low-mass secondary star, this implies that it is too large for its mass, possibly an evolved star that has lost a significant fraction of its envelope as a result of a long history of high thermal time-scale mass transfer. The thermal time-scale is of the order of $\tau_{\text{th}} \sim 10^8 (M_2/0.44 M_{\odot})^{-1.5} (R_2/1.5 R_{\odot})^{-2} \text{ yr}$ (e.g. Meintjes 2002, and references therein), which can be sustained for non-conservative mass transfer driven by mass and angular momentum losses from the system.

The possibility is also investigated whether the ~ 67 s pulsation could represent a beat period between a very short rotation period and the Keplerian period of denser X-ray emitting gas orbiting in the same direction as the WD, acting primarily as reflectors for an illuminating searchlight from the accreting WD. The corresponding spin-up time-scale for an $\sim 1.3 M_{\odot}$ WD accreting from a disc that extends to its surface is of the order of $\tau_{\text{s-u}} \leq 10^7 \text{ yr}$ (e.g. Wang 1987). This scenario is plausible, given the long thermal mass-transfer time-scale. If the ~ 67 s oscillation is a beat between the real spin period and hot reflecting blobs orbiting next to the accreting WD, one can readily show (e.g. Patterson 1979) that the spin period would then be

$$\frac{1}{P_*} = \frac{1}{P_b} + \frac{1}{P_K}, \quad (2)$$

with P_b representing the ~ 67 s beat period and P_K the Keplerian period of orbiting blobs close to the surface of the WD. Adopting the Hamada–Salpeter equation of state for an $\sim 1.3 M_{\odot}$ WD, the WD radius is $R_{\text{WD}} \approx 4 \times 10^8 (M_{\text{WD}}/1.3 M_{\odot})^{-0.8} \text{ cm}$ (Hamada & Salpeter 1961). Blobs orbiting the WD at distances r_K of the order of $R_{\text{WD}} \leq r_K \leq 10^9 \text{ cm}$ will have Keplerian periods of the order of $P_K \sim 4\text{--}15 \text{ s}$, resulting in WD rotation periods of the order of $P_* \sim 4\text{--}12 \text{ s}$ (frequencies $\sim 80\text{--}270 \text{ mHz}$), where the lower limit is slightly above the $P_{\text{b-u}} = 3.4 \text{ s}$ break-up period of the WD. However, no consistent oscillation was revealed between 80 and 270 mHz during the LS analysis of the higher frequencies.

The association of the consistent ~ 67 s with the spin period of the WD still leaves the question of the $\pm 3 \text{ s}$ drift observed in individual runs. It has been shown earlier (see also Fig. 8) that orbital motion cannot account for this drift. One possibility is an extended envelope around the accreting WD. In steady nuclear burning SSSs, a photospheric envelope with radius two to three times the radius of the accreting WD may exist (Ibragimov et al. 2003). For an extended dense hydrogen burning envelope around the WD, the spread in period can possibly be ascribed to the fact that the envelope is perhaps not quite synchronized (corotating) with the WD, resulting in a slippage and associated spread in the ~ 67 s oscillation.

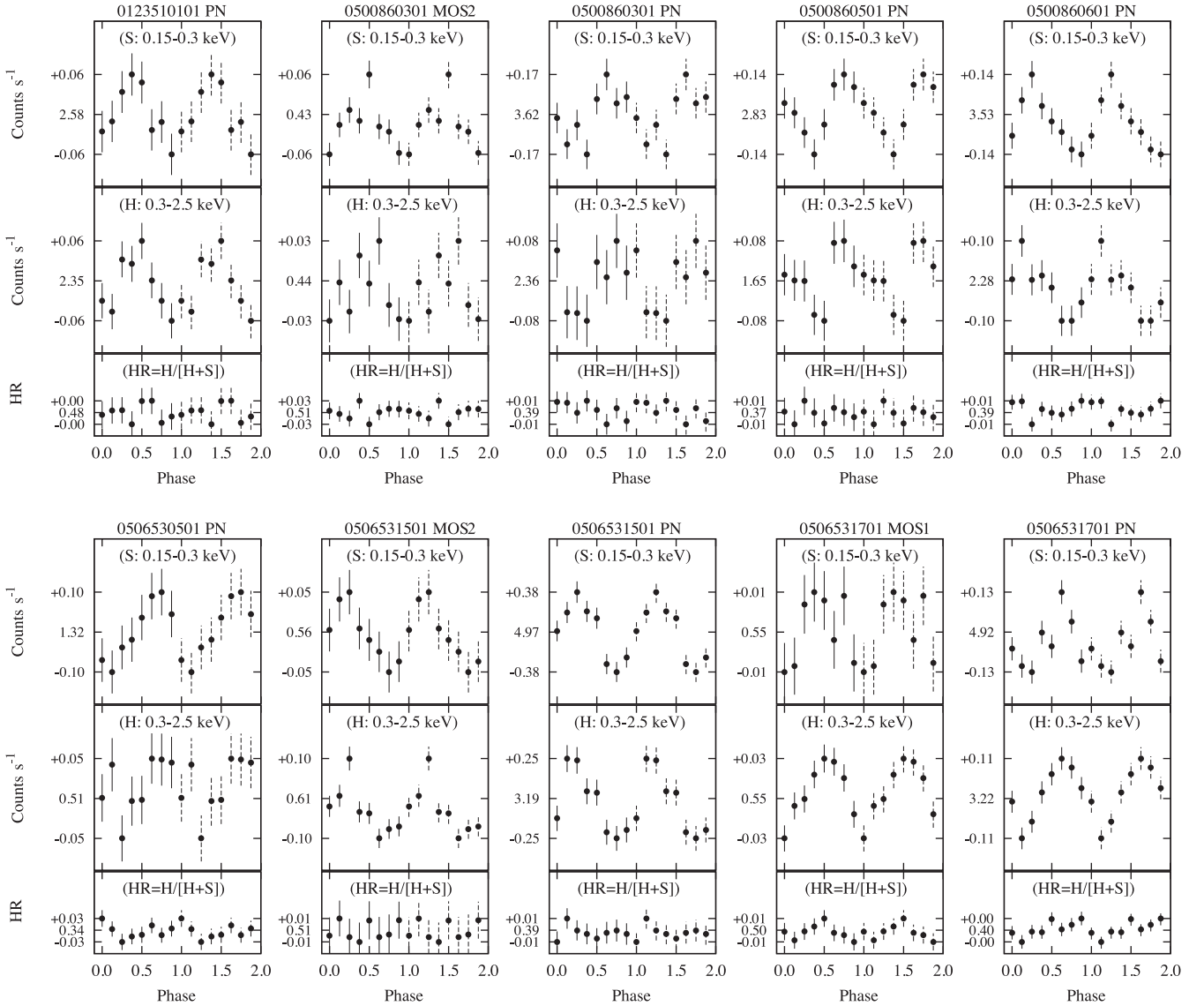


Figure 9. Soft (S), hard (H) and HR light curves of each data set folded on the corresponding period in Table 2, relative to the first point in each respective data set. The value halfway between the minimum and the maximum is indicated on the vertical axes, together with the semi-amplitude.

4 SUMMARY

A systematic search for short time-scale periodicities in archival *XMM-Newton* data of CAL 83 has revealed a consistent period around $\sim 67 \pm 3$ s. Stacking of the periodograms of all on-state observations, including those with no evidence of the pulse, showed that the ~ 67 s pulsation is present at a $\geq 3\sigma$ significance level. The presence of a consistent ~ 67 s oscillation may be an indicator of the rotation period of a highly spun-up accreting WD with an extended nuclear burning envelope not quite synchronized with the rotation period of the WD, hence the ± 3 s period drift around a median period of 67 s. Other, more extreme possibilities of a beat period between the rotation period of a WD spun up to a period just above the break-up period and nearby orbiting reflecting blobs could not be verified.

If the 67 s oscillation represents the spin period of a highly spun-up WD, it implies that CAL 83 possesses one of the fastest known rotating WDs in the compact binary population. The rapid rotation can possibly be explained in the context of a high thermal time-scale mass-transfer history of an evolved $\sim 0.44 M_{\odot}$ secondary star,

filling its $\sim 1.5 R_{\odot}$ Roche lobe. The presence of an oversized secondary relative to its mass may point towards an evolved secondary star that has shed the bulk of its envelope in a runaway mass-transfer history. Evidence for an evolved secondary star may be the presence of CNO cycling which can be inferred from the presence of carbon, oxygen and nitrogen lines in optical and UV spectra. This may place CAL 83 on a similar evolutionary path to the CV AE Aqr. The possibility of the association of the ~ 67 s oscillation with the rotation period of a WD with a nuclear burning envelope justifies detailed follow-up studies.

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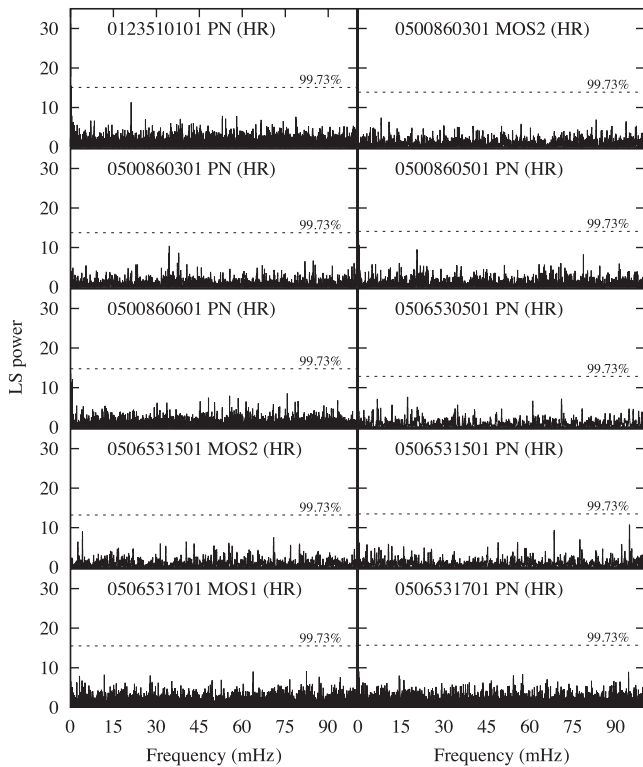


Figure 10. LS periodograms of HR light curves. The 99.73 per cent significance level is indicated, calculated over the 0–100 mHz range.

2.7.3 (www.python.org) was used for some of the data analysis in this work. This research has made use of NASA's Astrophysics Data System.

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Periodic X-ray modulations in Supersoft X-ray Sources¹

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Abstract. The supersoft source CAL 83 is often considered to be the prototype of its class. We report the discovery of modulations at a period of ~ 67 s in *XMM-Newton* X-ray data of CAL 83. This may be the spin period of a highly spun-up white dwarf, which is to be further investigated with follow-up observations in both the X-ray and optical wavebands. The supersoft source SMC 13 has an orbital period of ~ 4.1 h. SMC 13 was reported in the literature to exhibit orbital modulation in its X-ray flux, as inferred from its folded *ROSAT* light curve. We report the confirmation of this orbital modulation from three *Chandra* data sets, each providing continuous coverage of ~ 2.7 complete orbital cycles.

1. Introduction

Supersoft X-ray sources (SSS) form a highly luminous ($\sim 10^{36}$ - 10^{38} erg s⁻¹) class of objects that emit more than $\sim 90\%$ of their energy below 0.5 keV. SSS were first observed with the *Einstein X-ray Observatory* (Long et al. 1981, Seward and Mitchell 1981) and further discoveries by *ROSAT* established them as a distinct class of objects (Trümper et al. 1991). Typical effective temperatures range between ~ 15 -80 eV (see Kahabka and van den Heuvel (2006) for a recent review on SSS).

It was shown by van den Heuvel et al. (1992) that the low effective temperatures and high luminosities of these sources can be explained by the nuclear burning of hydrogen on the surface of a white dwarf accreting material on the thermal time scale of the donor. The accretion rate required for steady nuclear burning is $\sim 10^{-7} M_{\odot}$ yr⁻¹, which can be sustained if the donor mass is comparable to or greater than the white dwarf mass (Paczynski 1971, Savonije 1983). Many SSS are also believed to contain accretion discs.

Recent analysis of archived data of the binary supersoft sources CAL 83 and SMC 13 showed strong modulations in the X-ray waveband. We report preliminary results of the timing analysis performed on these data, based on the results obtained during the M.Sc. research of Odendaal (2012).

¹ Based on observations obtained with *XMM-Newton*, an ESA science mission with instruments and contributions directly funded by ESA Member States and NASA, as well as observations from the *Chandra Data Archive*.

2. CAL 83 in the Large Magellanic Cloud

2.1. Observations and data calibration

CAL 83 has been observed by *XMM-Newton* several times from April 2000 to May 2009. The X-ray data were recorded with the three EPIC detectors (MOS1, MOS2 and PN) and the RGS. The data sets were calibrated by following standard data reduction procedures with the *XMM-Newton* Science Analysis System², Version 11.0. The arrival times in the event files were corrected to the solar system barycentre (TDB system). From each calibrated event file, a light curve with a binning of 10 s was created from the source photons in the 0.15-2.5 keV range.

2.2. Timing analysis

Each of the CAL 83 X-ray light curves were detrended by subtracting the mean and dividing by the standard deviation. The task SCARGLE in the Starlink PERIOD package³ was subsequently utilized to create a Lomb-Scargle (LS) periodogram from each detrended light curve (Scargle 1982).

A power peak at a $\gtrsim 2\sigma$ level near 15 mHz ($P \sim 67$ s) was found in the EPIC PN (the most sensitive of the 5 detectors) periodograms of 6 of the observations. These periodograms are shown in figure 1. The EPIC MOS periodograms of some of these observations have a very weak feature near ~ 67 s. The periodicity was not detected in the RGS periodograms.

The EPIC PN data of the 6 data sets exhibiting the periodicity were reanalysed by subdividing the longer data sets into shorter light curves and creating a similar LS periodogram for each. The results from the segments exhibiting the periodicity are summarized in table 1. The error bars on the period values were determined by the intrinsic Fourier resolution of each periodogram, which gives an error of $P^2/2T$ to each side of each period peak, where P is the period, and T the total length of the light curve.

It is evident from table 1 that the detected period varies over a range of several seconds. The power of the peak is also highly variable for the different periodograms. No obvious correlation was found between the mean count rate and peak power, although it was noted that the period was not detected in data sets where the EPIC PN count rate was below 1 count/s.

2.3. Discussion

The exact origin of the ~ 67 s period is still under investigation. As the periodicity was found in the X-ray data, it is associated with the white dwarf. The fact that the modulation at ~ 67 s is present in different data sets spanning over about 9 years possibly rules out short-lived quasi-periodic oscillations. A very promising possibility is that the periodicity represents the spin period of a magnetized white dwarf in CAL 83. The period is reminiscent of the short white

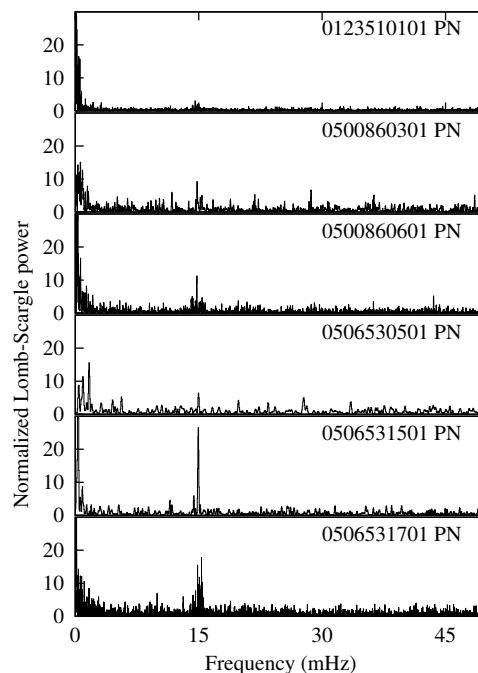


Figure 1. LS periodograms of EPIC PN observations exhibiting ~ 67 s periodicity.

² <http://xmm.esa.int/sas/>

³ <http://www.starlink.rl.ac.uk/star/docs/sun167.htx/sun167.html>

Table 1. Timing analysis of the CAL 83 *XMM-Newton* PN light curves exhibiting a ~ 67 s periodicity. “Peak power” refers to normalized power from Lomb-Scargle periodogram.

Obs ID	Length (s)	BJD (middle)	Period (s)	Peak power	Mean PN counts/s	Principal investigator
0123510101	5000	2451657.95459	68.6 ± 0.5	7.8	7.3	Fred Jansen
	5000	2451658.06917	67.9 ± 0.5	9.8	5.9	
	5000	2451658.12704	67.2 ± 0.5	5.2	6.7	
	5000	2451658.24162	67.2 ± 0.5	12.5	3.4	
	5000	2451658.29949	66.3 ± 0.4	7.7	3.0	
0500860301	5240	2454288.51696	65.0 ± 0.4	7.8	6.9	Thierry Lanz
	5240	2454288.57761	67.5 ± 0.4	8.6	6.8	
0500860601	5010	2454429.44780	67.8 ± 0.5	5.1	5.9	
	5010	2454429.56378	67.8 ± 0.5	9.3	6.9	
	5010	2454429.62176	69.9 ± 0.5	10.6	7.2	
0506530501	4600	2454573.12196	66.8 ± 0.5	6.6	2.1	Robert Schwarz
0506531501	6480	2454691.16206	66.9 ± 0.3	27.6	9.3	
0506531701	4570	2454981.86631	67.6 ± 0.5	7.8	8.8	
	4570	2454981.91920	66.6 ± 0.5	10.5	8.7	
	4570	2454981.97210	66.1 ± 0.5	15.9	9.2	
	4570	2454982.02499	64.9 ± 0.5	6.4	9.4	
	4570	2454982.07788	64.9 ± 0.5	5.5	7.9	
	4570	2454982.13078	67.3 ± 0.5	8.5	10.0	
	4570	2454982.18367	67.6 ± 0.5	4.7	9.8	
	4570	2454982.23657	66.1 ± 0.5	6.9	9.8	
	4570	2454982.28946	65.4 ± 0.5	15.6	8.9	
	4540	2454982.34224	65.3 ± 0.5	16.5	9.2	

dwarf spin periods in the cataclysmic variables AE Aqr (~ 33 s), V533 Her (63.633 s) and DQ Her (142 s) (Norton et al. 2004).

If it does represent the white dwarf spin period, one would expect the detected period to have exactly the same value in different data sets. However, it must be kept in mind that, although the arrival times have been corrected to the solar system barycentre, the orbital motion of the white dwarf will also modulate the detected spin period. This possibility is still under investigation.

Consistent modulations in other supersoft sources that have been ascribed to the spin period of a magnetized white dwarf are e.g. the 217.7 s pulsations in the source XMMU J004252.5+411540 in M31 (Trudolyubov and Friedhorsky 2008), and the 1110 s pulsations in Nova M31N 2007-12b (Pietsch et al. 2011). The recurrent nova RS Oph exhibited an unstable ~ 35 s periodicity during the supersoft phase after its 2006 outburst (see Osborne et al. (2011) and references therein). The favoured explanation of this variable modulation is the possibility of non-radial oscillations caused by nuclear burning instabilities. It may be that the ~ 67 s X-ray periodicity in CAL 83 is due to a similar effect.

3. SMC 13 in the Small Magellanic Cloud

The orbital period of SMC 13 was first reported by Schmidtke et al. (1994), but the most recent orbital ephemeris is that of van Teeseling et al. (1998) (with the number on the left denoting the time of minimum light, and the one on the right the orbital period of 0.1719260 ± 0.0000007 days):

$$T_0 = \text{HJD } 2450434.1320 \pm 0.0006 \quad + \quad 0.1719260E \pm 0.0000007 \text{ d}.$$

Kahabka (1995) reported the discovery of orbital modulation in the *ROSAT* data of SMC 13, and Kahabka (1996) determined the *ROSAT* period to be 4.123 h (0.1718 d).

Table 2. Archived *Chandra* X-ray observations of SMC 13.

Obs ID	Instrument	Start date & time (UT)	Exposure time (s)	PI
4535	ACIS-S3	2005-01-30 16:56:17	40 140	J. Greiner
7456	HRC-S+LETG	2007-02-12 18:25:16	40 190	T. Lanz
8519	HRC-S+LETG	2007-02-18 00:42:08	42 670	T. Lanz

3.1. Observations and data calibration

We discuss 3 archived *Chandra* X-ray *Observatory* observations of SMC 13, which are summarized in table 2. The calibration of the data was carried out by following standard data reduction and processing procedures with the CIAO software⁴, Version 4.3, using Version 4.4.5 of the CALDB (CIAO Calibration Database). The arrival times in the calibrated event files of the three observations were corrected to the solar system barycentre (TDB system).

3.2. Timing analysis

The PERIOD task SCARGLE was used to search for an orbital period in the *Chandra* X-ray light curves around the period of 4.123 h found by Kahabka (1996) in the *ROSAT* data. Each of the three Lomb-Scargle periodograms exhibited a strong power peak at the approximate position of the orbital period. However, the uncertainty in the peak positions was very high due to the relatively short length of the individual data sets. Therefore the X-ray modulation could be constrained to nothing better than 4.44 ± 0.89 h with the separate observations.

To obtain a higher period resolution, two additional Lomb-Scargle periodograms were created (see figure 2 for the region around the strongest peak): one by combining all three observations, and the other by combining only the observations of February 2007 (Obs 7456 and 8519). The presence of numerous alias peaks is evident, as well as possible harmonics of the orbital period. The strongest peak in the 2007 periodogram was determined to be at a period of 4.12 ± 0.06 h. The position of this (relatively broad) peak was then used to choose the appropriate peak on the high time-resolution periodogram obtained by combining all three data sets, yielding a period of 4.1214 ± 0.0005 h. The error value was calculated with the method described in Section 2.2, considering the total time from the start of Obs 4535 to the end of Obs 8519.

The *ROSAT* light curves of Kahabka (1996) in the soft (S: 0.1-0.25 keV) and hard (H: 0.26-0.50 keV) energy bands, as well as the hardness ratio, folded on the ephemeris of Schmidtke et al. (1994) are shown in figure 3.

The intrinsic energy resolution of the ACIS-S detector was used to create two similar *Chandra* light curves: one in the soft band (S: 0.100-0.250 keV) and one in the hard band (H: 0.251-0.500 keV). For both the *ROSAT* and the *Chandra* light curves, the hardness ratio was calculated as $HR = (H-S)/(H+S)$. The soft, hard and HR light curves of observation 4535 were folded on the newly determined period by making use of the task **efold** in the Xronos Timing Analysis Software Package⁵, using the time of minimum light of the third minimum of the Obs 4535 light curve (BJD 2453401.5682), and these are provided in figure 4.

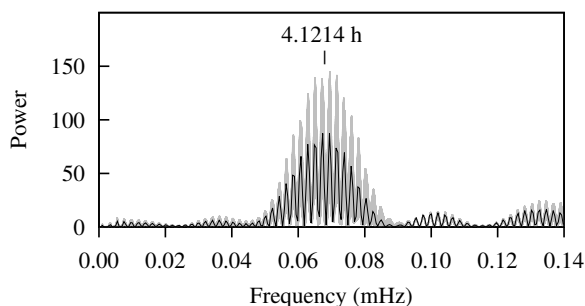


Figure 2. Lomb-Scargle periodograms of combined SMC 13 *Chandra* light curves. Black: Obs 7456 and 8519. Gray: Obs 4535, 7456 and 8519.

⁴ <http://cxc.harvard.edu/ciao/index.html>

⁵ See the Xronos User's Guide (HEASARC 2009) for more information.

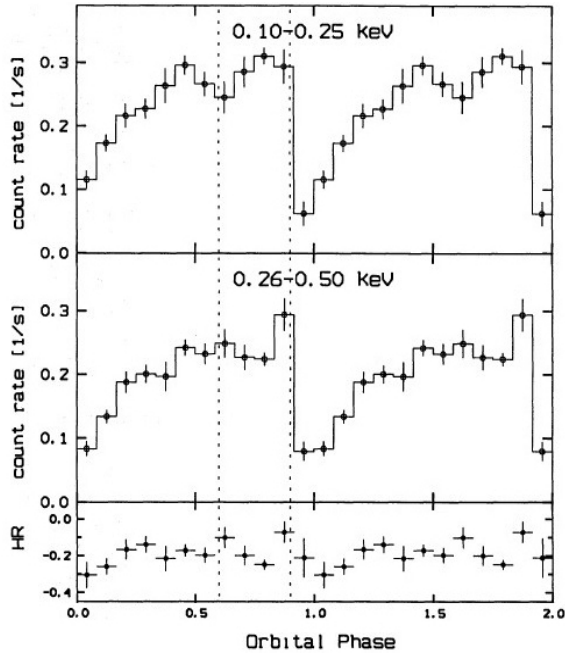


Figure 3. *ROSAT* light curves of SMC 13, folded on 4.123 h with respect to the epoch Nov. 3.105, 1994. (Adopted from Kahabka (1996, figure 3).)

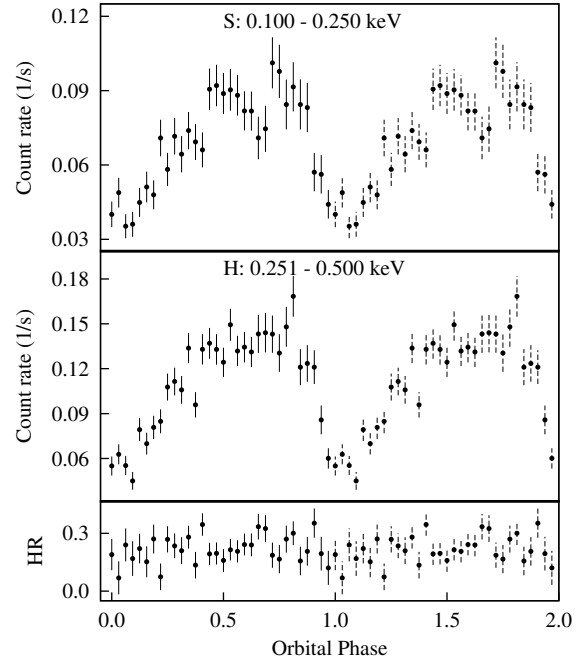


Figure 4. *Chandra* ACIS light curves of SMC 13, folded on $P_{\text{orb}} = 4.1214$ h with respect to the third minimum in the data set.

3.3. Discussion

Because the time of minimum in the *ROSAT* light curve is largely defined by only three data points, Crampton et al. (1997) mentioned the possibility that these few minima might be caused by the system being in a low state at some epochs, rather than by orbital modulation. However, each of the *Chandra* data sets constitutes an uninterrupted observation of the source for a duration of ~ 2.7 full orbital periods, and the orbital modulation is clearly evident.

The orbital period of 4.1214 ± 0.0005 h determined from the combined data set is very close to the $P_{\text{orb}} = 4.126224 \pm 0.000017$ h determined by van Teeseling et al. (1998) from photometrical data, although not included in its error ranges. However, the aliasing effects arising from the long time gaps between the *Chandra* observations made it very difficult to choose the appropriate peak, therefore the van Teeseling et al. (1998) period is considered to be more reliable.

Comparison of the folded *ROSAT* and *Chandra* light curves in figure 3 and figure 4 shows that they exhibit the same sharp, asymmetric form. As pointed out by Kahabka (1996), there seems to be a second dip at $\phi = 0.6$ in the 0.1-0.25 keV (S) band in addition to the main minimum at $\phi \sim 0.9-1.1$ in the *ROSAT* data, which is not present in the 0.26-0.50 keV (H) band. This is also visible in the *Chandra* light curves. The large amplitude variation may be ascribed to the eclipse of the primary taking place in a high inclination system. If this is the case, the fact that the eclipse is not total indicates that there is an extended emitting region. The structure of the light curve may be the result of variable absorption in the system, possibly by an accretion disc edge. As mentioned below, Kahabka (1996) also considered the possibility of SMC 13 being a polar-like system.

The hardness ratios (HR) for the 2 detectors can not be compared directly, as the H and S counts depend on detector sensitivity. However, there does seem to be an increase in HR during the minor dip, and just before the major dip for both *ROSAT* and *Chandra*. It is also interesting to note that the count rate in the H band correlates with the variations in HR. The favoured

explanation of Kahabka (1996) for this HR modulation is that it may be due to temperature variations, as an increase in temperature would result in an increase in both the HR and the count rate. Kahabka (1996) suggested that a changing observed temperature can be explained if the white dwarf is magnetized, as this would result in a changing viewing angle of the polar caps or an accretion column at the polar caps.

From figure 3, it appears that the X-ray modulation is in phase with the orbital motion. On the other hand, Schmidtke et al. (1996) found that the X-ray minimum occurs about 0.25 of an orbital cycle earlier than the optical minimum, according to their ephemeris. However, folding the *Chandra* light curves through the orbital ephemeris of van Teeseling et al. (1998) showed that the X-ray minima approximately coincides with the optical minimum. Therefore a more in-depth analysis of the orbital period and ephemeris in conjunction with previous observations of SMC 13 needs to be carried out. In particular, simultaneous optical and X-ray observations are needed to clarify the relative phasing of the minima in these wavebands. If the minima they do not coincide, it can indicate a situation where the main sources of X-ray and optical flux respectively are not aligned and therefore not eclipsed simultaneously, for example with the X-rays being brightest close to the compact object, and an optically bright “hot spot” where the accreting stream hit an accretion disc rim.

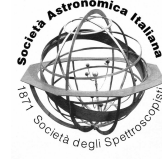
Probing periodicities in supersoft sources can therefore be instrumental in understanding these systems. With timing analysis, orbital and spin periods and also various types of quasi-periodic signals can be detected and can provide information on the orbital parameters and component masses, as well as the nature of the compact object and accretion process.

Acknowledgments

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X-ray periodicity in the supersoft X-ray sources CAL 83 and SMC 13

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Abstract. The discovery of X-ray modulations in CAL 83 with a period of ~ 67 s is reported. Its long-term presence suggests it may be the spin period of a highly spun-up white dwarf; however, this is still under investigation. SMC 13 has an orbital period of ~ 4.1 h, and has been reported in the literature to exhibit orbital modulation in its ROSAT X-ray flux. This orbital modulation is confirmed from 3 *Chandra* observations, each providing continuous coverage of ~ 2.7 orbital cycles.

Key words. X-rays: binaries – Stars: individual: CAL 83 – Stars: oscillations – Stars: individual: SMC 13 – binaries: eclipsing

1. Introduction

Supersoft X-ray sources (SSS) form a highly luminous ($\sim 10^{36} - 10^{38}$ erg s⁻¹) class of objects that emit $\gtrsim 90\%$ of their energy below 0.5 keV. Van den Heuvel et al. (1992) showed that this can be explained by the nuclear burning of hydrogen on the surface of a white dwarf (WD) accreting at a very high rate ($\sim 10^{-7} M_{\odot}$ yr⁻¹) from a companion star.

2. CAL 83 in the LMC

To date, the binary SSS CAL 83 has been observed 24 times by XMM-Newton. These archival observations were calibrated, and barycentric corrections performed, with SAS

v. 12.0.1. A Lomb-Scargle (LS) periodogram was created from each light curve with the Starlink PERIOD package v. 5.0-2.

A peak near 15 mHz (~ 67 s) was discovered at a 6.3σ significance level in the periodogram of the PN detector (the most sensitive of the X-ray detectors) of observation 0506531501, and at a lower significance level in 6 other PN data sets. In 3 of these 7 observations, one of the MOS light curves also exhibited a weak peak at ~ 67 s (see Table 1). The periodogram and light curve for observation 0506531501 are shown in Fig. 1.

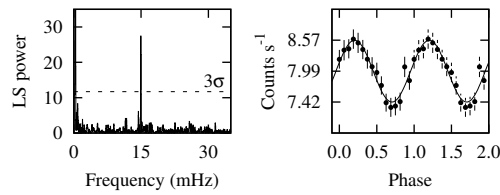
The long-term presence of the oscillation leads to its interpretation as the WD spin period. However, the ~ 2 s variation from the mean in the period, not associated with the orbital motion, complicates this straightforward interpretation, and is still being investigated.

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Table 1. Results of Lomb-Scargle analysis of CAL 83 XMM-Newton light curves.

Obs ID	Start date	Period (s)	σ
0123510101PN	23-04-2000	68.63 ± 0.06	$^{-1}$
0500860301MOS2	06-07-2007	66.1 ± 0.2	1.5
0500860301PN		67.5 ± 0.2	2.1
0500860501PN	05-10-2007	70.6 ± 0.2	$^{-1}$
0500860601PN	24-11-2007	67.7 ± 0.1	2.5
0506530501PN	16-04-2008	66.8 ± 0.5	1.0
0506531501MOS2	12-08-2008	66.8 ± 0.3	2.8
0506531501PN		66.9 ± 0.3	6.3
0506531701MOS1	30-05-2009	67.29 ± 0.05	2.8
0506531701PN		65.17 ± 0.05	4.3

¹Faintly visible, but statistically insignificant peak.

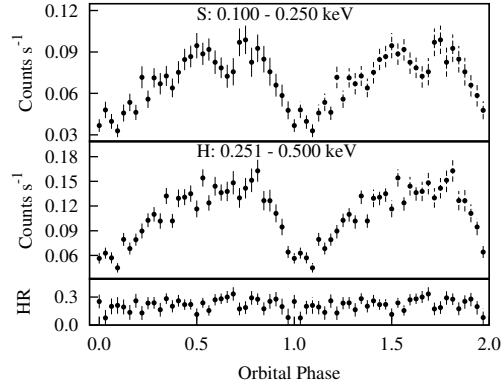
**Fig. 1.** Left: LS periodogram of data set 0506531501 PN. Right: Folded light curve. A sine curve with a period of 66.9 s is overplotted.**Table 2.** *Chandra* observations of SMC 13.

Obs ID	Instrument	Start date	Exposure (s)
4535	ACIS-S3	2005-01-30	40140
7456	HRC-S+LETG	2007-02-12	40190
8519	HRC-S+LETG	2007-02-18	42670

3. SMC 13 in the SMC

Kahabka (1996) reported the discovery of orbital modulation with a period of 4.123 h in the ROSAT X-ray light curve of the binary SSS SMC 13. According to Van Teeseling et al. (1998), the orbital period is 4.126 h.

Timing analysis was performed on 3 archival *Chandra* X-ray Observatory observations of SMC 13 (see Table 2), each constituting an uninterrupted observation of ~ 11 h. Calibrated light curves were obtained with CIAO v. 4.3, and an LS periodogram was created for each. The orbital modulation was evident, and the periodograms exhibited substantial power in the region of 4.123 h. All 3 observations were then combined to obtain

**Fig. 2.** *Chandra* ACIS light curves of SMC 13, folded on $P_{\text{orb}} = 4.1378$ h with respect to third minimum of data set. $\text{HR} = (\text{H}-\text{S})/(\text{H}+\text{S})$.

a periodogram with a higher period resolution. Aliasing effects made it difficult to decide which peak represents the true orbital period, so the period of Van Teeseling et al. (1998) was used to choose the most appropriate peak, yielding $P_{\text{orb}} = 4.1378 \pm 0.0005$ h.

The intrinsic energy resolution of ACIS-S was used to create soft, hard and hardness ratio (HR) light curves (Fig. 2). Comparison with the ROSAT light curve of Kahabka (1996) shows remarkable stability over more than a decade. The form of the light curves may be due to SMC 13 being an eclipsing system, where the observed X-ray emission is modulated by the accretion disc edge. The X-ray eclipse is not total, indicating that we are observing the modulation of an extended region.

Acknowledgements. The financial assistance of the South African SKA Project towards this research is hereby acknowledged. Opinions expressed and conclusions arrived at are those of the authors. We would like to thank the PI's for the opportunity to analyse the archived data, as well as Frank Haberl for his comments on the CAL 83 results.

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CAL 83 — The Prototypical Close Binary Supersoft X-ray Source in the LMC: A Short Review

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Abstract

CAL 83 is a luminous supersoft X-ray binary that is often considered the prototype of its class. It is believed to contain an accreting white dwarf with surface hydrogen burning, and has an orbital period of ~ 1 day. X-ray spectroscopy indicates a massive white dwarf with luminosity $\sim 10^{37}$ erg s⁻¹ and effective temperature ~ 46 eV. The rich optical spectrum provides evidence of a precessing accretion disc with an associated outflow, as well as a surrounding ionization nebula. The system exhibits X-ray to optical variability on time-scales of $\gtrsim 1$ min to > 1 year. A brief review of this intriguing binary in the Large Magellanic Cloud is presented.

Keywords: stars: individual: CAL 83 - binaries: close - X-rays: binaries - accretion, accretion discs - white dwarfs - stars: oscillations - stars: winds, outflows

1 Introduction

Supersoft sources (SSS) are extremely soft, high-luminosity X-ray emitters, with typical effective temperatures of a few hundred thousand Kelvin, and bolometric luminosities of $L_{\text{bol}} \sim 10^{36}$ - 10^{38} erg s⁻¹, close to the Eddington limit for a solar-mass object (e.g. [21]). The first luminous SSS were discovered in the Large and Small Magellanic Clouds (LMC and SMC) with the *Einstein Observatory* [26, 43], and further observations during the *ROSAT* all-sky survey [47] and subsequent pointed observations established SSS as a new class of objects.

CAL 83 was the brightest X-ray point source discovered in the Columbia Astrophysics Laboratory *Einstein* survey that is not associated with a Galactic foreground star or a background AGN, with a luminosity of a few times 10^{36} erg s⁻¹ in the 0.15-4.5 keV band. A blue object with magnitude $B \sim 16.8$ was identified as the optical counterpart [5]. Due to the similarities of its optical properties to that of low-mass X-ray binaries, this was initially the favoured model for CAL 83, i.e. a system containing an accreting neutron star or black hole [8, 22, 23, 32, 44].

However, Van den Heuvel et al. [48] proposed instead that many SSS consist of a white dwarf (WD) in a binary system, with the energy released by the nuclear burning of accreted hydrogen on the WD surface responsible for the high X-ray luminosity. The accretion rate required to sustain nuclear burning is very high ($\sim 10^{-7}$ M_⊙ yr⁻¹), and can be sustained if the donor is more massive than the WD.

Photometrical observations showed the orbital period to be around 1 day [7, 8, 42, 44]. The most recent value was obtained from the analysis of OGLE III data spanning 2001-2009, i.e. 1.047529 ± 0.000001 d [36]. The folded light curve is quasi-sinusoidal in shape, with no observed eclipses, indicating a relatively low binary inclination.

During the past two and a half decades since its discovery, observational studies of CAL 83 have proven it to be a very fascinating source, the exact nature of which is everything but clear-cut. In this paper, a short review is presented, focussing on its properties in the X-ray to optical wavebands. Little is known at longer wavelengths: In the near-infrared, weak H emission and only a marginal detection of the Ca II infrared triplet were reported [7]. ATCA radio observations indicated an upper limit of < 0.12 mJy at wavelengths of 3.5 cm and 6.3 cm [11].

2 The X-ray spectrum

The supersoft nature of the X-ray emission from CAL 83 was confirmed by observations with *EXOSAT* in 1983 and 1985 [8, 32], *ROSAT* in 1990-1991 [14, 41, 47], *BeppoSAX* [34] and *ASCA* [10].

Initially, blackbody models provided satisfactory fits to the low-resolution X-ray spectra. However, the absorbing neutral hydrogen column density N_{H} along the line of sight was uncertain, as

it would include interstellar absorption not only in the Galactic foreground, but also in the LMC, as well as any intrinsic absorption in the binary itself. Blackbody fits with different N_{H} yielded very different values for kT_{eff} and L_{bol} .

Modelling of the Ly α absorption in the *HST* UV spectrum indicated that $N_{\text{H}} = (6.5 \pm 1.0) \times 10^{20} \text{ cm}^{-2}$ along the line of sight to CAL 83, which is consistent with the best fit value of $\sim 7 \times 10^{20} \text{ cm}^{-2}$ obtained from a WD model atmosphere analysis of the *ROSAT* spectrum [12]. This value is in good agreement with the Galactic foreground hydrogen absorption, indicating that CAL 83 must be on the near side of the LMC, and that no highly significant neutral hydrogen absorption occurs within the source itself.

Later it came to light that a blackbody fit tends to overestimate the WD luminosity, because actual WD atmospheres emit a larger fraction of their total energy in the soft X-ray band than a blackbody with the same temperature. Therefore a WD model atmosphere fit to an X-ray spectrum with a certain flux will translate to a lower value of the integrated bolometric luminosity than a blackbody fit, and this will also lead to an underestimation of the WD mass. Also, after the first high-resolution spectra ($\sim 0.05 \text{ \AA}$) of CAL 83 obtained with *XMM-Newton* [31] and *Chandra* became available, it was obvious that the fine structure of the spectrum, including various absorption features, can not be fit by a smooth continuum spectrum to a high degree of accuracy.

Various authors performed fits of WD atmosphere models with varying degrees of sophistication to the low-resolution CAL 83 spectra [12, 17, 20, 34, 46]. LTE models assume local thermodynamic equilibrium, while the NLTE (non-LTE) models also incorporate departures from LTE.

The most recent investigation involved detailed NLTE model atmosphere analysis of the first high-resolution X-ray spectra of CAL 83: an *XMM-Newton* RGS spectrum (23 April 2000, obs ID 0123510101) in combination with a *Chandra* LETG spectrum (15 August 2001, obs ID 1900) [25]. The best fit parameters when fixing N_{H} to $6.5 \times 10^{20} \text{ cm}^{-2}$ are $\log g = 8.5 \pm 0.1$, where g is the WD surface gravity, $kT_{\text{eff}} = 46 \pm 2 \text{ eV}$, photospheric radius $R = (6.96 \pm 0.70) \times 10^8 \text{ cm}$ and $L_{\text{bol}} = (3.4 \pm 1.1) \times 10^{37} \text{ erg s}^{-1}$. These values correspond to a WD mass in the $1.0\text{--}1.4 M_{\odot}$ range, and this was the first direct spectroscopic evidence that the WD in CAL 83 is massive.

It should be noted that the layers of accreted, hydrogen-burning material on the surface of a WD in a SSS presents conditions quite different from that on the surface of a WD without accretion. In fact, an accreting, hot WD with steady surface nuclear burning will have a photospheric radius 2 to 3 times larger than that of a cool WD [17].

3 Anti-correlated long-term X-ray and optical variability

It was originally thought that CAL 83 is a persistent X-ray source, but it has since been observed during an X-ray off-state on the following dates: 28 April 1996 (*ROSAT*, [18, 19]), 29-30 November 1999 (*Chandra*, [13]), 3-4 October 2001 (*Chandra*, [25]), 2 January - 18 March 2008 (*Swift*, [15, 24]), 16 January, 10 March, 11 April and 16 April 2008 (*XMM-Newton*, [36]).

Different models explaining the X-ray off-states in terms of a decrease [1] or increase [20] in the accretion rate have been proposed. The 1992-1999 MACHO light curve exhibited variability of $\lesssim 1$ mag, and upon correlation with the epochs of the X-ray on- and off-state observations, it was found that the observed X-ray off-states were all during an optical high state, while the X-ray on-states were during a lower optical state [13].

This anti-correlation was confirmed by comparing the OGLE light curve and the more recent (2007-2009) *XMM-Newton* observations [36]. The long-term optical variability was found to be quasi-periodic ($P \sim 450 \text{ d}$), and the X-ray on (optical low) state has a length of $\sim 200 \text{ d}$. This implies $M_{\text{WD}} \sim 1.32\text{--}1.38 M_{\odot}$. Applying blackbody fits to the *XMM-Newton* spectra indicated a slightly higher X-ray temperature during the X-ray high states than during the X-ray low states.

The anti-correlation can be explained by a qualitative model where an increase in the accretion rate causes an increase in the WD photospheric radius, lowering kT_{eff} and shifting the emission peak towards the extreme UV, causing an X-ray off-state and increasing the optical luminosity [13, 36, and references therein]. The irradiation of the disc increases, also enhancing the optical luminosity. After the accretion rate decreases again, the WD photosphere contracts and the source enters an X-ray high, optical low state again. Possible reasons for a modulation in the mass transfer rate may be: (i) attenuation when a magnetic star spot on the secondary passes over the L_1 point, or (ii) the attenuation of the mass transfer from the secondary by a strong wind from the X-ray

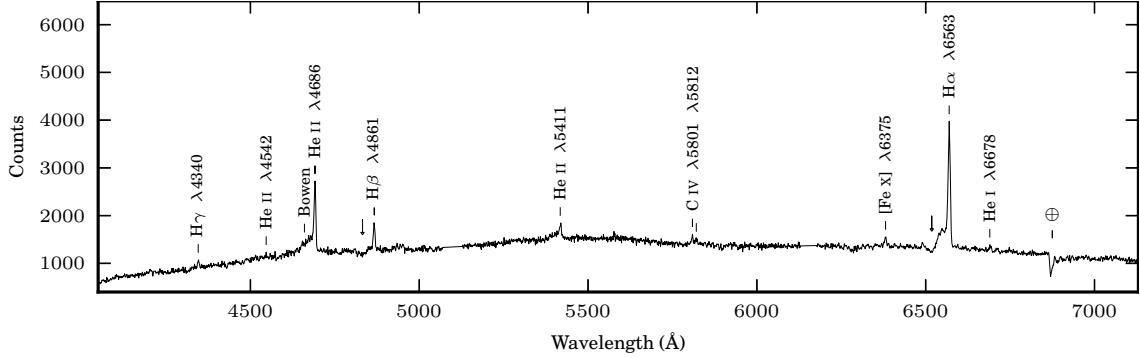


Figure 1: SALT RSS spectrum of CAL 83 on 5 January 2013. The resolution is ~ 3.8 Å. Arrows indicate the H α and H β P Cyg absorption features. The $\oplus$ indicates the atmospheric B-band.

source stripping off the outer layers of the secondary.

4 The optical and UV spectrum

4.1 The central source

The optical spectrum of CAL 83 contains emission lines typical of an accretion disc. The most prominent features are the Balmer and He II emission lines [3, 6, 8, 9, 16, 30, 32, 44]. These lines have variable broad wing structures extending to several thousand km s^{-1} from the main component. The wings are sometimes observed to be blueshifted, and sometimes redshifted relative to the main component. A wind or weakly collimated jet from an accretion disc precessing with a period of ~ 69 d could be responsible for the variable wing structures [8]. Although this is not yet well established, it is in good agreement with theoretical predictions [45]. P Cyg profiles in the Balmer lines provide additional support for the outflow scenario. A spectrum obtained with the Robert Stobie Spectrograph (RSS) on the Southern African Large Telescope (SALT) is shown in Fig. 1, exhibiting blueshifted wing structures¹.

Unfortunately, the donor star is too faint to establish its spectral type. Several of the papers above determined the radial velocity semi-amplitude of the He II $\lambda 4686$ emission line to be in the 30–40 km s^{-1} range, leading to the conclusion that the secondary must be significantly less massive than the WD, which contradicts the model of Van den Heuvel et al. [48]. However, this mass function determination assumes that the He II line provides an exact measure of the WD orbital motion, which would only be true if the distribution of its emission was perfectly symmetrical about the WD. A stellar wind is expected to blow from the heated side of the donor, and Van den Heuvel et al. [48] argued that the He II $\lambda 4686$ emission line originates from the interaction between this wind and a wind from the WD/disc, and that its radial velocity semi-amplitude is much smaller than the actual value for the WD.

The optical spectrum also contains evidence of ionized carbon and nitrogen, including the Bowen blend at ~ 4660 Å, and also He I $\lambda 6678$, [Fe X] $\lambda 6375$ and O VI emission lines at 3811, 3834 and 5290 Å. The ratio of hydrogen emission to He II and CNO emission is relatively low [8]. This serves as evidence of CNO cycling in the donor star, indicating an evolved donor of which the outer layers have already been stripped by accretion and possibly winds from the X-ray source.

Ultraviolet spectra obtained with the *IUE* during 1984–1987 showed emission lines of N V $\lambda 1240$, Si IV $\lambda 1403$ and He II $\lambda 1640$, as well as several interstellar absorption features [2, 8]. In January 1987, the UV flux was higher by a factor of 2 to 3, and ESO observations showed that this coincided with an optical high state. The shape of the UV/optical continuum indicated emission from an accretion disc rather than pure stellar emission. Modelling of the spectral energy distribution also showed that the observed optical and UV flux can be explained very well by a combination of the

¹Proposal ID: 2012-2-RSA.UKSC-006. PI: A. Odendaal

WD spectrum, the accretion disc with the reprocessing of radiation and disc flaring included, and also some emission of reprocessed radiation from the heated side of the donor star [35].

From the *HST* GHRS spectrum, the N v $\lambda 1239$ and $\lambda 1243$ resonance doublet lines were determined to have a FWHM of $\sim 1 \text{ \AA}$, and red wings extending to $\sim +800 \text{ km s}^{-1}$ [12]. These lines are expected to originate from the disc, and the small widths indicate that the orbital inclination is relatively low. The red wing structures may have the same origin as the blue- or redshifted wing of the He II $\lambda 4686$ emission.

Eleven observations with *FUSE* showed that the UV flux is modulated by the orbital period, and that the radial velocity semi-amplitude of the O VI $\lambda 1032$ emission line is $23 \pm 6 \text{ km s}^{-1}$, similar to that of He II $\lambda 4686$ [42]. However, the relative phasing of these lines are not the same, therefore they probably originate in different parts of the system. They also do not have the same phasing as the photometric orbital light curve.

4.2 The ionization nebula

From a theoretical investigation, it was predicted that SSS should be surrounded by ionized regions, with the central X-ray emitter as the ionizing source [37]. Evidence of forbidden oxygen emission from an extended region was indeed found in the optical spectra, and an [O III] $\lambda 5007$ image confirmed the presence of an ionized nebula around CAL 83 [33]. Detailed spectroscopic studies showed [O III], [O I], [N II], [Si II], He II $\lambda 4686$ and Balmer emission from a nebula with a dense inner region ($\sim 5\text{--}10 \text{ cm}^{-3}$) with a typical radius of $\sim 7.5 \text{ pc}$ and a more diffuse outer region extending to $\sim 20 \text{ pc}$ [16, 38].

The mass contained in the inner nebula was estimated to be $\sim 150 M_{\odot}$, while the material ejected from the binary system itself (e.g. by an accretion wind) should only be of the order of $\sim 1 M_{\odot}$. Therefore the nebula is primarily caused by the ionization of the local ISM by the supersoft X-rays from CAL 83. Calculations of the time-dependent properties of ionization nebulae around SSS indicated that CAL 83 has been radiating at a time-averaged luminosity of $\gtrsim 3 \times 10^{37} \text{ erg s}^{-1}$ for at least the past $\sim 10^5$ years, with any luminosity variations taking place on time-scales $\lesssim 2 \times 10^3 \text{ yr}$ [4].

5 Short time-scale variability

5.1 Optical magnitudes

Variability of a few tenths of a magnitude on time-scales of 2 h or less were reported [8, 32, 44]. We performed differential photometry with the Sutherland High-speed Optical Camera (SHOC) on the SAAO 1.9-m Telescope, and our Lomb-Scargle (LS) analysis revealed significant peaks in the $\lesssim 1 \text{ mHz}$ region in some of the periodograms [28]. Their positions are not constant and they have broad profiles. We ascribe these to quasi-periodic oscillations associated with the accretion disc around the WD.

5.2 X-ray modulations on time-scales of several minutes

A 38.4 min pulsation was discovered in *Chandra* observation 1900, which was not present in *XMM-Newton* observation 0123510101 [40]. It was ascribed to non-radial g-mode pulsations in the WD, similar to those that have recently been observed in some novae. We performed a LS analysis on the 18 more recent *XMM-Newton* observations (proposal IDs 050086 and 050653), and found variable peaks above a 99.73% significance level in 13 of the light curves, ranging from ~ 10 to >100 minutes [27]. Five of the periodograms have a peak at the same position (within error bars) as the 38.4 min pulsation, and several different peaks are often present in the same periodogram. These transient periods may also be ascribed to non-radial g-mode pulsation in the WD, driven by the ϵ -mechanism, i.e. instabilities caused by thermonuclear reactions (e.g. [39]), but more detailed modelling is needed to explain their variable nature.

We also quantified the correlation between the count rate and hardness ratio for all the light curves by making use of the Spearman linear rank-order correlation coefficient, and established a significant correlation ($>99.73\%$) for 4 observations, indicating that higher count rates are associated with a higher source temperature even on short time-scales (e.g. Fig. 2)[27].

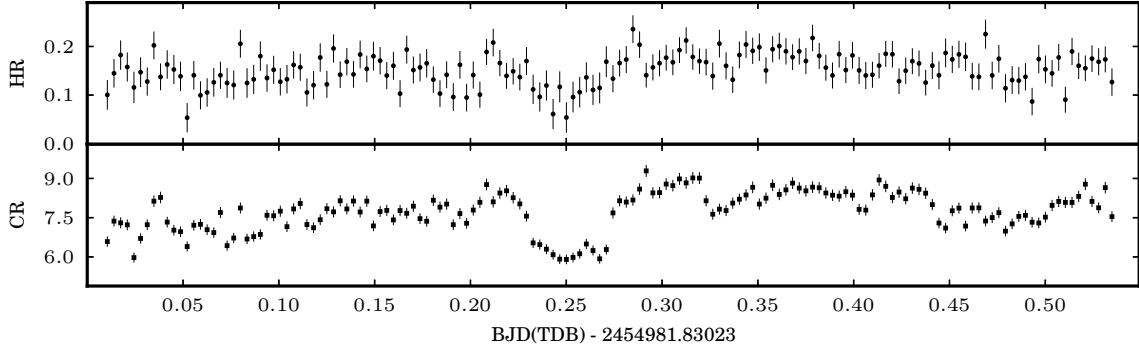


Figure 2: EPIC pn hardness ratio (top) and count rate (bottom) light curves of *XMM-Newton* observation 0506531701 of CAL 83.

5.3 The ~ 67 s pulsation

In 7 of the 19 on-state *XMM-Newton* observations, a ~ 67 s X-ray pulsation was discovered [29]. There is significant variability in the period, exhibiting a spread of up to ~ 3 s from the median of ~ 67 s. If this period is related to the WD rotation, its variability needs to be explained. One possibility is that we may be observing the WD rotation through an extended H-burning envelope, with these layers “slipping” on the fast rotating WD surface, with their rotation not quite synchronized with the WD spin. An alternative interpretation is that this pulsation is also related to non-radial oscillations of the WD.

6 Summary

Observational evidence indicates that CAL 83 contains a massive WD that has already passed through a long period of mass accretion from a companion star, of which the mass has not been determined conclusively. It appears that the mass transfer rate from the companion undergoes cyclic modulations, causing the quasi-periodic superorbital modulation in the optical magnitude that is anti-correlated with the long-term changes in the X-ray flux. While the optical modulation is well defined by the MACHO and OGLE light curves, X-ray monitoring (e.g. with *Swift*) is needed to better define the shape of the long-term X-ray light curve.

It has already been proposed that these superorbital modulations are related to temperature changes, but a correlation between X-ray flux and hardness ratio can also be found on time-scales of minutes to hours. X-ray oscillations on time-scales of a few minutes and more are probably associated with non-radial WD pulsations. This possibility needs to be explored by modelling the pulsation modes that can be expected from the accreting WD.

The optical and UV properties are consistent with emission from an accretion disc, with a wind or a jet with a large collimation angle originating from the disc. This outflow may be radiation-driven, but if the WD has a short spin period, the ejection of accreting blobs from the fast rotating WD magnetosphere is also a possibility. Obtaining optical spectra spread over a few months should shed more light on the possible jet precession.

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DISCUSSION

JIM BEALL: Is the Doppler shifted line you see evidence for disk rotation or more likely from a jet?

ALIDA ODENDAAL: The broadened main components of the emission lines indicate emission from an accretion disc, albeit with some contribution from the heated side of the secondary. However, the additional broad wing structures and P Cyg features are displaced from the main components by several thousand km s^{-1} , comparable to the WD escape velocity. Therefore the Doppler shifted wings that we see are most probably the result of a jet with a large collimation angle from the inner accretion disc.

DMITRY BISIHALO: Do you have any information about the value of the magnetic field in the system?

ALIDA ODENDAAL: No, there has not been a direct measurement of the WD magnetic field strength, e.g. by means of polarimetry. However, if the 67 s pulsation represents the WD spin period, then (i) the detection of the spin period indicates inhomogeneous emission from the WD surface, perhaps due to the accretion flow being channelled onto the polar caps by the magnetic field, and (ii) it indicates that the WD has been strongly spun up by disc torques during its lifetime, which can take place much more effectively in the presence of a magnetic field.



Multiwavelength variability in CAL 83[★]

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Abstract. CAL 83 is a supersoft X-ray source in the LMC that exhibits multiwavelength variability on a wide range of time-scales. The optical magnitude shows quasi-periodic variability with a period of ~ 450 d, an orbital modulation of just more than a day, and also possible QPOs associated with an accretion disc. The X-ray lightcurves exhibit large-amplitude modulations, including periodicities ranging from 67 s to hours. The X-ray variability is accompanied by temperature changes, and is anti-correlated with the optical behaviour.

Key words. Stars: individual: CAL 83 — X-rays: binaries — White dwarfs — Stars: oscillations — Accretion, accretion discs

1. Introduction

Supersoft X-ray sources (SSS) are highly luminous, with bolometric luminosities of the order $\sim 10^{37}$ erg s⁻¹. They are “supersoft” because their X-ray emission peaks at low-energy X-rays, typically between ~ 15 and 80 eV (see e.g. the review of Kahabka & van den Heuvel 2006). The X-ray emission observed in SSS can be explained by nuclear hydrogen burning on the surface of an accreting white dwarf (WD) (Van den Heuvel et al. 1992).

As illustrated in Fig. 2 of Hachisu & Kato (2001), the nature of the H burning on the WD surface is determined by the WD mass M_{WD} and the accretion rate $\dot{m}_{\text{acc}}$. For material consisting of 70% H, $\dot{m}_{\text{acc}} \sim (3.7\text{--}7.6) \times 10^{-7} (M_{\text{WD}}/M_{\odot} - 0.40) M_{\odot} \text{ yr}^{-1}$ can sustain steady H shell burning.

For lower $\dot{m}_{\text{acc}}$, steady burning is not possible, but shell flashes trigger nova outbursts. For $\dot{m}_{\text{acc}}$ above the steady burning regime, H burning can take place in a thin shell around the WD, while excess material piles up to form an extended envelope, with part of the envelope blown off in an optically thick wind.

CAL 83 is a close binary supersoft source in the LMC. It has long been considered the prototype of persistent SSS, but is in fact highly variable. In this paper, its known multiwavelength variability is summarized, and new results based on *XMM-Newton* data are presented.

2. Long-term variability

CAL 83 exhibits pronounced brightness changes (~ 1 mag) in the optical. These modulations are quasi-periodic with a period of ~ 450 d (Rajoelimanana et al. 2013, hereafter R13). Since its discovery, CAL 83 has been observed 8 times during X-ray off-states, co-

[★] Based on observations obtained with *XMM-Newton*, an ESA science mission with instruments and contributions directly funded by ESA Member States and NASA.

inciding with optically high states. The X-ray luminosity during the on-state observations is also variable. R13 presented new results on this anti-correlated optical and X-ray variability, as well as a review of previously published results.

The anti-correlation between the optical and X-ray flux may be caused by changes in $\dot{m}_{\text{acc}}$. A small increase in $\dot{m}_{\text{acc}}$ will cause excess material to pile up, increasing the WD photospheric radius, as well as the irradiation of the accretion disc and the rate of inward flow through the disc. The shift of the WD spectrum towards longer wavelengths, as well as the increased irradiation and resulting emission of reprocessed radiation from the disc (and the disc wind?) will cause a decrease in the X-ray luminosity and an increase in the optical luminosity (also see Greiner & Di Stefano 2002 and references therein). Once the excess material has been drained, or the mass transfer rate from the secondary decreases, the contraction in the WD photosphere shifts the emission peak to soft X-rays again.

3. Other optical variations

The orbital modulation in the optical lightcurve was refined by R13 to 1.047529(1) d (with a ~ 0.1 mag semi-amplitude).

Photometrical observations with the Sutherland High-Speed Optical Camera (SHOC) on the SAAO 1.9-m Telescope in April 2013 revealed quasi-periodic modulations at frequencies $\lesssim 1$ mHz (as summarized in Odendaal et al. 2013). They are most probably quasi-periodic oscillations (QPOs) from the disc.

Optical spectra reveal broad wing structures in the He II and Balmer emission lines, as well as P Cyg profiles in the Balmer lines. These structures are variable, and probably originate in a wind/jet from the disc, which is possibly precessing at a period of ~ 69 d (Crampton et al. 1987).

4. The ~ 67 s X-ray pulsation

The *XMM-Newton* archive contains 19 on-state and 4 off-state observations of CAL 83. In 7

of the on-state observations, we discovered a ~ 67 s X-ray pulsation (Odendaal et al. 2014). However, it is not completely stable, exhibiting a spread of up to ~ 3 s from the median.

If this is related to the WD rotation, then its variability needs to be explained. We may be observing the WD rotation through an extended envelope surrounding it. The rotation of the nuclear burning envelope may not be quite synchronized with the WD spin, with these layers “slipping” on the fast rotating WD surface, modulating the observed period.

Alternatively, the pulsation may be associated with g-mode oscillations, driven by the ϵ -mechanism, i.e. due to an instability caused by thermonuclear reactions (e.g. Saio 2013).

5. The 38.4 min X-ray pulsation

From a NLTE analysis of *Chandra* observation 1900 and *XMM-Newton* observation 0123510101 of CAL 83, Lanz et al. (2005) obtained $M_{\text{WD}} \sim 1.3 \pm 0.3 M_{\odot}$ and a photospheric radius 2.5 times that of the degenerate core.

Schmidtke & Cowley (2006) (hereafter SC06) performed a detailed timing analysis of these 2 observations, and discovered a period of ~ 38.4 min in the *Chandra* lightcurve. They ascribed this periodicity to non-radial pulsations of the accreting WD.

Non-radial g-mode WD pulsations first became known in single WDs, where they are detected in the optical and UV (e.g. Córscico 2009). They are often multi-periodic and often stable on time-scales of years. Non-radial pulsations have also been observed in dwarf novae (e.g. Warner & Woudt 2005).

More recently, pulsations on time-scales of several minutes have also been observed in X-rays from accreting WDs with surface H burning. Drake et al. (2003) found a strong periodic signal of ~ 41.7 min in the nova V1494 Aql, together with a suite of periods from 8.8 to 57.7 min. Ness et al. (2003) reported a ~ 22.1 min pulsation in V4743 Sgr, with two weaker harmonic overtone periods at 11.1 and 7.5 min. The 38.4 min pulsation in CAL 83 may be of a similar origin.

Since the paper of SC06, another 18 on-state observations of CAL 83 were obtained

with *XMM-Newton*. We took another close look at these archival observations to investigate variability on time-scales of several minutes to hours.

6. Recent *XMM-Newton* lightcurves

The observations were calibrated with the SAS Version 13.0.1. Broadband (0.15–2.5 keV) EPIC lightcurves with barycentric corrections were extracted. Soft (0.15–0.25 keV), hard (0.25–2.5 keV) and hardness ratio lightcurves were also created. The hardness ratio was defined as $HR = (H-S)/(H+S)$, with S and H the soft and hard count rates respectively.

6.1. Hardness ratio correlations

The broadband and HR lightcurves are shown in Fig. 1 for the EPIC pn detector. The lightcurves exhibit significant variability. It is also evident that the count rate (CR) and hardness ratio (HR) seem to follow the same trends for many of the observations.

To quantify this correlation, the Spearman linear rank-order correlation coefficient r_s was utilized, (Press et al. 1986, p. 749). r_s describes the correlation between two simultaneous lightcurves and can vary between -1 (complete anti-correlation) and $+1$ (complete correlation), with 0 indicating no correlation. The Spearman coefficients correlating the CR and HR for each observation in Fig. 1 respectively were calculated, and are listed in Table 1 together with their statistical significances.

A correlation between CR and HR was confirmed above a 99.73% significance level for observations 0500860401, 0500860501, 0500860601 and 0506531701, and at a lower level in some of the other observations. The only negative r_s values were obtained for 0506531201 and 0506531601, where the count rates were extremely low, but the significance levels of these anti-correlations are also low.

Table 1 also gives the average pn CR and average pn HR for each observation, and these are plotted in Fig. 2. Here one can also see a trend according to which observations with high average CRs also have high average HRs, which agrees with the findings of R13. For this

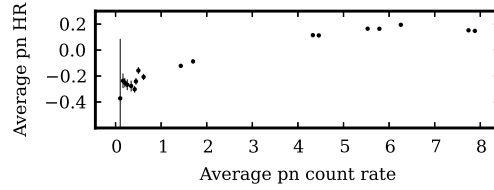


Fig. 2. The average EPIC pn hardness ratio versus the average EPIC pn count rate for the CAL 83 observations listed in Table 1.

correlation, $r_s=0.909$, with a 99.999983% significance. We can conclude that, in general, an increase in the X-ray count rate is accompanied by an increase in temperature. It has already been proposed that the superorbital modulations (§2) are related to temperature changes, but it now appears that a correlation between X-ray flux and temperature can also be found on time-scales of minutes to hours.

6.2. Lomb-Scargle analysis

For each observation, a combined EPIC lightcurve in the 0.15–2.5 keV range was created by obtaining a weighted average of the mos1, mos2 and pn lightcurves. Lightcurves with a wide range of bin sizes were created, and detrending was performed by subtracting different low-order polynomial functions.

A comprehensive Lomb-Scargle (LS) analysis (with the Starlink PERIOD package¹) was performed to search for periodic behaviour in the X-ray lightcurves at periods longer than the known ~ 67 s (15 mHz) pulsation. Many of the observations had significant peaks at frequencies below 2 mHz. The statistical significance levels were determined by performing Monte Carlo significance tests. The LS periodograms for the observations exhibiting periodic behaviour are shown in Fig. 3. These periodograms were produced using a bin size of 50 s, with long term trends removed by subtracting a second-order polynomial fit.

The peaks in Fig. 3 that have a $>99.73\%$ significance level are marked, and listed in Table 2. It should be noted that the indicated peaks were also visible for the other bin sizes

¹ www.starlink.ac.uk/docs/sun167.htm/sun167.html

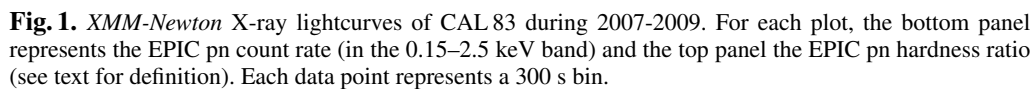


Table 1. *XMM-Newton* observations of CAL 83 from 2007 to 2009. The average EPIC pn count rate and hardness ratio is given for each observation. The Spearman correlation coefficient r_s between the pn count rate and hardness ratio lightcurves (see Fig. 1) is also given, together with its significance level.

Observation	Start date & time (UT)	Average pn CR	Average pn HR	Spearman coefficient r_s	Significance level (%)
0500860201	2007-05-13 22:03:32	6.257 ± 0.031	0.195 ± 0.005	0.320	94.9713
0500860301	2007-07-06 23:31:52	5.786 ± 0.031	0.164 ± 0.006	0.487	99.7000
0500860401	2007-08-21 15:11:21	4.457 ± 0.032	0.113 ± 0.007	0.580	99.7627
0500860501	2007-10-05 23:49:21	4.327 ± 0.024	0.115 ± 0.006	0.548	99.9826
0500860601	2007-11-24 21:10:14	5.529 ± 0.022	0.165 ± 0.004	0.598	99.99989
0500860901	2008-03-21 02:10:24	0.204 ± 0.006	-0.251 ± 0.036	0.019	9.3808
0506530201	2008-03-20 00:33:22	0.615 ± 0.013	-0.207 ± 0.024	0.292	83.4068
0506530501	2008-04-16 12:32:01	1.699 ± 0.026	-0.087 ± 0.016	0.168	45.0144
0506530601	2008-04-17 13:40:13	1.431 ± 0.015	-0.121 ± 0.011	0.227	81.6993
0506530801	2008-04-19 06:43:35	0.342 ± 0.012	-0.277 ± 0.041	0.323	80.8932
0506530901	2008-04-20 22:38:43	0.419 ± 0.009	-0.302 ± 0.026	0.121	52.4060
0506531001	2008-04-21 18:47:57	0.500 ± 0.012	-0.157 ± 0.026	0.356	94.2082
0506531201	2008-04-23 11:20:22	0.163 ± 0.008	-0.237 ± 0.056	-0.463	97.4067
0506531301	2008-04-25 08:13:32	0.448 ± 0.010	-0.241 ± 0.026	0.221	76.0228
0506531401	2008-04-29 00:40:23	0.257 ± 0.006	-0.267 ± 0.041	0.189	78.7418
0506531501	2008-08-12 14:50:27	7.881 ± 0.046	0.148 ± 0.006	0.371	90.2642
0506531601	2008-09-17 11:10:21	0.102 ± 0.007	-0.372 ± 0.458	-0.194	59.9341
0506531701	2009-05-30 08:00:48	7.741 ± 0.017	0.152 ± 0.002	0.536	>99.99999

Table 2. Periodicities detected at a >99.73% significance level in the periodograms (Fig. 3) of the combined EPIC lightcurves.

Observation	Pulsation periods (in minutes)
0500860201	68.2 $\pm$ 12.5; 37.5 $\pm$ 3.8
0500860401	37.2 $\pm$ 5.8 ; 13.1 $\pm$ 0.7
0500860501	136.7 $\pm$ 46.9; 37.3 $\pm$ 3.4
0500860601	81.3 $\pm$ 10.3; 52.0 $\pm$ 4.2
0500860901	24.3 $\pm$ 1.5
0506530201	94.0 $\pm$ 39.2
0506530501	9.8 $\pm$ 0.7
0506530601	140.7 $\pm$ 58.6; 35.2 $\pm$ 3.5
0506531301	119.3 $\pm$ 49.7; 29.8 $\pm$ 3.0; 22.9 $\pm$ 1.8
0506531401	112.1 $\pm$ 28.5
0506531501	46.3 $\pm$ 10.4
0506531601	82.0 $\pm$ 34.2; 19.5 $\pm$ 1.9
0506531701	303.0 $\pm$ 61.2; 151.5 $\pm$ 15.2; 116.5 $\pm$ 9.0

and detrending variants, and there were no significant peaks in the background periodograms at these positions.

It is interesting to note that 5 of the periodograms has a peak at the same position (within error bars) as the 38.4 min reported by

SC06. There are also peaks at other periods, and often several simultaneous peaks. We think that these transient periodicities are also related to non-radial g-mode pulsations as discussed in §5, possibly driven by nuclear burning instabilities. The multiple peaks may represent harmonic periods. These are preliminary results and will be investigated in more detail. We also need to explain the wide range of and variability in the observed pulsations. We believe, though, that the unstable and ever-changing accreted layers on the WD surface in CAL 83 can be expected to yield a much more variable series of pulsations to be triggered than in WD systems without H burning or accretion.

7. Conclusions

The optical and X-ray variability of CAL 83 on various time-scales has been reviewed. The count rates and hardness ratios of the recent *XMM-Newton* X-ray lightcurves have been explored, and it was found that the significant modulations in the count rates are often correlated with the hardness ratio variations. The X-ray modulations also exhibit periodic be-

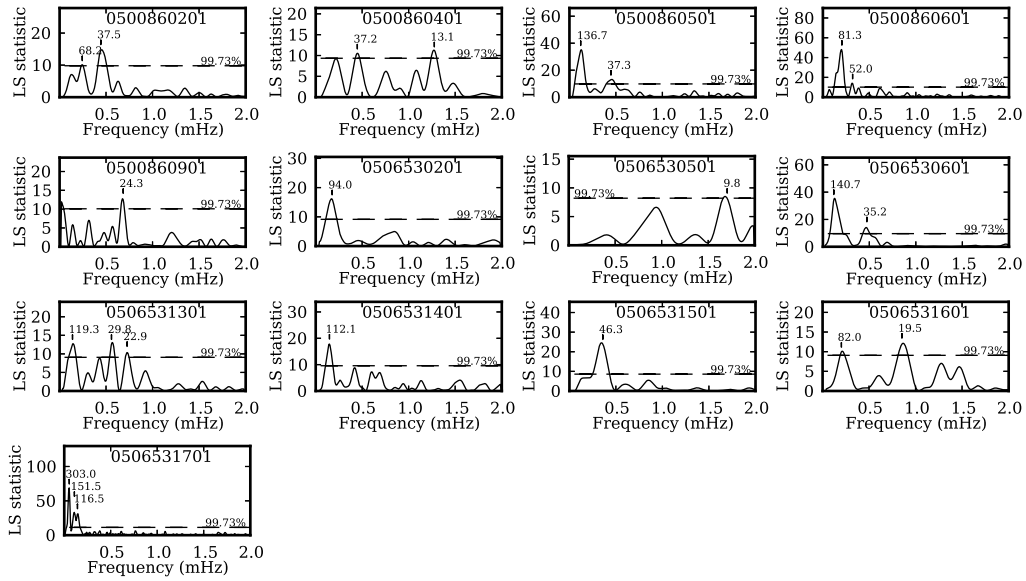


Fig. 3. Lomb-Scargle periodograms of those combined EPIC lightcurves that exhibit significant peaks. The lightcurves had a binning of 50 s and were detrended by subtracting a 2nd-order polynomial fit. The 99.73% significance level is indicated. Each significant peak has been marked with its period in minutes.

haviour, with pulsation periods ranging from ~10-300 minutes. The possibility that these pulsations are associated with non-radial WD pulsations is currently being explored.

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Supersoft X-ray Source CAL 83: A Possible AE Aqr-like System

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Abstract

CAL 83 is a close binary supersoft X-ray source in the Large Magellanic Cloud. A ~ 67 s periodicity detected in supersoft X-rays is most probably associated with the spin period of a highly spun-up white dwarf (WD). The variability in the period is ascribed to the obscuration of the WD by the hydrogen burning envelope surrounding it, rotating with a period that is close to, but not quite synchronized with, the WD rotation period. Optical spectra obtained with SALT exhibit accretion disc emission lines with broad wing structures and P Cyg profiles, indicating mass outflows. Timing analysis of photometrical observations performed at the South African Astronomical Observatory (SAAO) revealed variable signals at ≤ 1 mHz which are thought to be associated with quasi-periodic oscillations from an accretion disc. The short spin period inferred for CAL 83 can be the result of spin-up by accretion disc torques during a long mass transfer history, placing this source on a similar evolutionary track as the cataclysmic variable AE Aqr.

Keywords: stars: individual: CAL 83 - binaries: close - white dwarfs - stars: oscillations - stars: winds, outflows - stars: evolution.

1 Introduction

Supersoft X-ray sources (SSS) are highly luminous in the supersoft X-ray band, with $>90\%$ of the unabsorbed X-ray flux being below ~ 0.5 keV. These sources were first discovered by the *Einstein X-ray Observatory* and *ROSAT* (see Kahabka & van den Heuvel (2006) and references therein).

Van den Heuvel et al. (1992) (hereafter vdH92) showed that the low effective temperatures and high luminosities of SSS can be explained by the nuclear burning of hydrogen on the surface of a white dwarf (WD) accreting from a binary companion (the close binary SSS model). The accretion rate required for steady nuclear burning is $\sim 1 - 4 \times 10^{-7} M_{\odot} \text{ yr}^{-1}$. Such a high accretion rate can be sustained if the companion mass is comparable to or greater than the WD mass, as this will cause the Roche lobe of the donor to shrink and drive mass transfer on the thermal time-scale of the donor.

CAL 83 is a close binary SSS in the Large Magellanic Cloud. However, it is not a persistent X-ray source and several X-ray off-states have been observed, with long-term optical variability anti-correlated with the long-term X-ray variability, as described by e.g. Rajoelimanana et al. (2013). These authors also derived a refined orbital period of 1.047529(1) d for CAL 83 from OGLE-III photometry. The inclination has been estimated to be in the

$i = 20 - 30^{\circ}$ range (see Cowley et al. (1998) and references therein, hereafter Co98). Lanz et al. (2005) used NLTE models during a combined analysis of *XMM-Newton* and *Chandra* data, and their results indicate a massive WD ($M_{\text{WD}} \sim 1.3 \pm 0.3 M_{\odot}$). Schmidtke & Cowley (2006) found a periodic signal of 38.4-minutes in the *Chandra* LETG data of CAL 83, which they ascribed to possible non-radial pulsations in the accreting WD.

The optical spectrum of CAL 83 is characterized by Balmer and He II accretion disc emission lines (Co98 and references therein). The He II $\lambda 4686$ emission line has a broad, variable wing structure that has also been observed in some of the Balmer lines. In each particular spectrum, these wing structures are either all towards the blue side of the main components of these emission lines, or all towards the red side. According to Crampton et al. (1987) (hereafter Cr87), this may be associated with the slow precession of an accretion disc with outflows through either a wind or a weakly collimated jet.

The ~ 67 s X-ray periodicity is briefly summarized in §2.1 (a detailed discussion is provided in Odendaal et al. 2014). Preliminary results of the analysis of optical data are presented in §2.2 and §2.3. After this discussion, the possible evolutionary scenario of CAL 83 is brought into context with that of the CV AE Aqr in §3, followed by the conclusions in §4.

2 CAL 83: Observations and results

2.1 The ~ 67 s X-ray pulsation

The *XMM-Newton* archive contains 23 observations of CAL 83, 4 of which were during an X-ray off-state. A systematic search for short time-scale periodicities that may be associated with the WD spin was performed on the 19 on-state light curves. The Starlink PERIOD¹ package was used to obtain a Lomb-Scargle (LS) periodogram of each light curve.

A power peak at a period of 66.8 ± 0.4 s was discovered at a $> 99.9999\%$ significance level in the periodogram of the EPIC PN light curve of observation 0506531501, and at a lower significance level in 6 of the other PN periodograms. In 3 of these observations, 1 of the MOS light curves also exhibited a weak ~ 67 s periodicity. The LS periodogram of observation 0506531501, as well as the light curve folded on the 66.8 s period, are shown in Fig. 1.

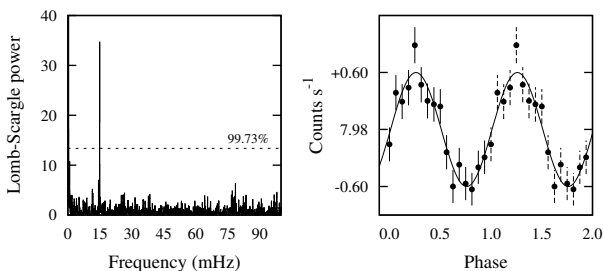


Figure 1: Lomb-Scargle periodogram (left) and folded light curve (right) of CAL 83 *XMM-Newton* dataset 0506531501 PN. Both the periodogram and folded light curve display a significant 66.8 s oscillation.

The overall significance of this detection, considering the PN datasets of all 19 the on-state observations, is $\geq 3\sigma$. The detected period exhibits a spread of ± 3 s around ~ 67 s between different observations, and even within a single observation. This variability can not be explained by Doppler shifts due to the WD orbital motion.

The long-term presence of the pulsation suggests its association with the WD spin, although the variation in the detected period complicates matters. It is possible, however, that we are observing the WD spin, but through an extended envelope around the accreting WD. According to Ibragimov et al. (2003), the photospheric radius in a steady nuclear burning SSS may have a radius 2-3 times the radius of the WD itself. The spread in the observed period can possibly be explained by the envelope not being quite synchronized with the WD, resulting in a slippage effect and associated variation in the pulsation period.

¹www.starlink.rl.ac.uk/star/docs/sun167.htx/sun167.html

2.2 Optical spectroscopy with SALT

A series of optical spectra of CAL 83 has been obtained with the Robert Stobie Spectrograph (RSS) on the Southern African Large Telescope (SALT), using the pg0900 grating to obtain a resolution of $R \sim 1500$. These spectra show similar characteristics to those discussed in §1, with strong Balmer and He II emission lines. The velocity widths of these lines indicate emission from regions compatible with an accretion disc around the WD. These lines also exhibit blue wing structures extending to ~ 2000 km s⁻¹ from the line centres, supporting the existence of a wind or weakly collimated jet in the system. P Cyg profiles are also evident in the H α and H β lines, serving as further evidence of mass outflow. Shown in Fig. 2 are the H α and H β profiles from a spectrum obtained on 2 February 2013.

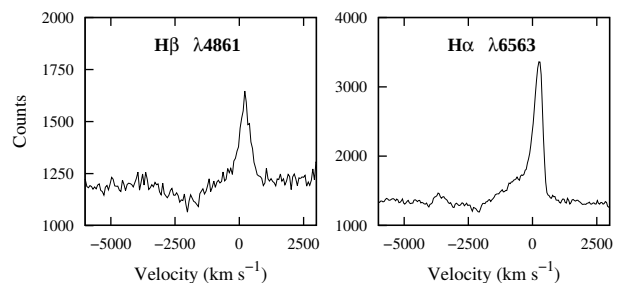


Figure 2: H β and H α profiles of CAL 83, obtained with SALT on 2 Feb 2013.

For $M_{\text{WD}} \sim 1.3 M_{\odot}$, the WD radius is $R_{\text{WD}} \approx 4 \times 10^8 (M_{\text{WD}}/1.3 M_{\odot})^{-0.8}$ cm (see Eracleous & Horne 1996 and references therein). The mass function can be determined if the semi-amplitude (K_1) of a spectral line originating close to the WD is known. Although the orbital coverage of the SALT spectra has not been adequate to calculate a new K_1 , various values have been reported in the literature. Adopting $K_1 \sim 35$ km s⁻¹, $i \sim 25^\circ$ (see Co98) and $P_{\text{orb}} = 1.047529$ d, the mass function is $4.7 \times 10^{-3} M_{\odot}$, and a secondary mass of $M_2 \sim 0.61 M_{\odot}$ is obtained, yielding a mass ratio $q = M_2/M_{\text{WD}} \sim 0.47$. However, this value of K_1 was derived from the radial velocity modulation of the He II $\lambda 4686$ disc emission line, and this line would only represent the WD orbital motion if the He II emission was distributed symmetrically about the WD. As discussed by e.g. Kaitchuck et al. (1994), there are some problems with this assumption, compromising the mass function calculation.

2.3 Fast photometry with SHOC

In April 2013, fast photometrical observations were obtained with one of the Sutherland High-speed Optical Cameras (SHOC2) on the SAAO 1.9-m Tele-

scope at Sutherland. These cameras have been commissioned recently, and are optimized for high time resolution photometry. A clear filter was used for the CAL 83 observations, and with this sensitive instrument, exposure times in the 0.5-5 s range were adequate, depending on observing conditions.

Aperture photometry was utilized to obtain light curves, and the CAL 83 light curves were corrected by performing differential photometry using 4 comparison stars. The Starlink PERIOD package was used to perform a LS analysis. The LS periodograms of CAL 83 did not reveal any significant pulsation at the inferred spin period of ~ 67 s or at higher frequencies.

However, some of the CAL 83 periodograms exhibited a significant peak in the region just below 1 mHz. These peaks have broadened profiles, and their positions do not stay constant. In Fig. 3, the LS periodogram of the SHOC observation on 15 April 2013 is presented, showing a peak at 1168 ± 214 s. The light curve folded on this period is also shown. This observation had a total length of 3245 s, and the light curve was binned to 20 s for this analysis. These peaks are ascribed to quasi-periodic oscillations associated with the accretion disc around the WD.

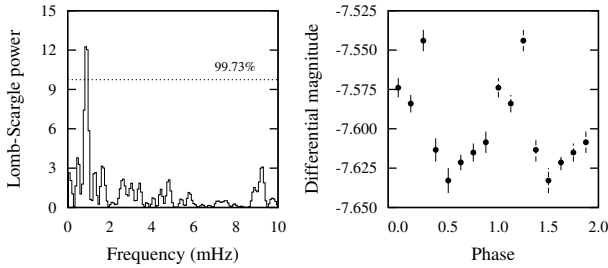


Figure 3: SHOC periodogram and light curve folded on the detected period of 1168 ± 214 s (obtained on 15 April 2013).

These quasi-periods at ≤ 1 mHz may be the Keplerian periods of blobs of material orbiting the WD in the accretion disc, or possibly beat periods between the presumed spin period of $P_* \sim 67$ s and blobs orbiting in the inner regions of the disc. Assuming that the 1168 ± 214 s periodicity is a Keplerian period, the associated Keplerian radius is $46 R_{WD}$. On the other hand, it may be a beat period between $P_* \sim 67$ s and a Keplerian period of ~ 71 s associated with blobs orbiting at $\sim 7 R_{WD}$. Adopting the orbital parameters in §2.2 yields a circularization radius of $\sim 150 R_{WD}$, supporting the compatibility of both the Keplerian radii above with emission from the accretion disc.

3 The evolution of CAL 83: an AE Aqr analogue?

AE Aqr is a nova-like variable with a short WD spin period (33.08 s), and a long orbital period (9.88 h). A

review of the multi-wavelength properties of AE Aqr is provided by Meintjes, Odendaal and van Heerden elsewhere in this volume. It has been shown by Meintjes (2002) and Schenker et al. (2002) that the properties of this source indicate a high mass transfer history, during which AE Aqr could have been a supersoft source. During this phase, accretion disc torques would have been able to spin-up the WD to such a short rotation period.

The probable association of the ~ 67 s period in CAL 83 with a short WD spin may place this source on a similar evolutionary path as AE Aqr. Evidence indicating that CAL 83 has already passed through quite a long mass transfer history does indeed exist. The high WD mass, just below the Chandrasekhar limit, is one of the factors supporting an extended period of mass accretion. However, with the WD being more massive than the WD in AE Aqr, it may be driven over the Chandrasekhar limit before entering the CV stage.

Emission lines of ionized carbon, nitrogen and oxygen are present in the optical and UV spectra of CAL 83, and as remarked by Cr87, the low ratio of H to He II and CNO lines indicate a donor with an H-poor envelope. Therefore the secondary may already have shed most of its envelope during a long period of mass transfer to the WD, with CNO cycling resulting in the observed line ratios. Assuming that the secondary mass is as low as $\sim 0.61 M_\odot$ (see §2.2), the Roche lobe of the secondary should be $R_2 \sim 1.7 R_\odot$. As the secondary has to be in contact with its Roche lobe for Roche lobe overflow, this implies a donor that is too large for its mass, possibly an evolved star that has lost a large fraction of its envelope, also supporting the long mass transfer history.

The thermal time-scale is of the order $\tau_{th} \leq 5 \times 10^7 (M_2/0.61 M_\odot)^{-2} (R_2/1.7 R_\odot)^{-1}$ yr (e.g. Meintjes 2002 and references therein), while the spin-up time-scale corresponding to a $\sim 1.3 M_\odot$ WD accreting from a disc extending to its surface is $\tau_{s-u} \leq 10^7$ yr (e.g. Wang 1987). These time-scales are comparable, showing that the WD in CAL 83 could have been spun-up to a short spin period if it has been in the supersoft phase with its characteristic high mass transfer rates for a sufficiently long time.

To sustain a $\tau \sim 10^7$ yr mass transfer, $q > 5/6$ is required for a conservative process, as described by the vdH92 model. However, evidence indicating significant mass and accompanied angular momentum losses is present in the optical spectra, suggesting a non-conservative mass transfer process. Given the possible WD spin period of ~ 67 s, the outflow mechanism may be the ejection of accreting material by a similar magnetospheric propeller mechanism as in AE Aqr. For non-conservative mass transfer, the Roche lobe will shrink according to the relation

$$\frac{\dot{R}_{L,2}}{R_{L,2}} = -\frac{\dot{M}_2}{M_2} \left[\frac{5}{3} - 2(1-\alpha)q - \frac{2}{3} \frac{q\alpha}{1+q} - 2\eta \frac{M_2 (GM_{WD} R_{\text{circ}})^{1/2}}{J_{\text{orb}}} \right] + \frac{2\dot{J}}{J}, \quad (1)$$

where α and η represent the fraction of mass and angular momentum transferred through L_1 that is lost from the system (e.g. Meintjes 2002). Therefore, for CAL 83, mass transfer that can sustain nuclear burning is perhaps only possible in the phase when the evolved donor is expanding, with significant angular momentum loss contributing substantially in keeping the Roche lobe in contact with the donor. This will be explored more quantitatively in the future.

4 Conclusion

It has been shown that CAL 83 exhibits a transient ~ 67 s X-ray periodicity. This is expected to be associated with the WD spin period, which would make it one of the shortest known spin periods in the white dwarf binary population. Results from optical data support the scenario of accretion through an accretion disc, creating the possibility of the WD being spun-up by disc torques during an extended period of mass accretion, similar to what happened during the evolution of AE Aqr. Optical spectra also present evidence of mass outflows in CAL 83, which can be related to the ejection of accreting blobs from a fast rotating WD magnetosphere or accretion disc winds. However, many questions about this fascinating source still remain unanswered, and detailed follow-up studies are essential.

Acknowledgement

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DISCUSSION

MARIKO KATO: How do you determine the secondary mass for CAL 83? Is it really so small? I expect $q \geq 1$ to explain anti-correlation with X-ray.

ALIDA ODENDAAL: The radial velocity semi-amplitude of the He II $\lambda 4686$ line was used to calculate $M_2 \sim 0.61 M_\odot$, which is only valid if the He II emission is uniformly distributed around the WD. This assumption can be highly problematic (e.g. Kaitchuck et al. 1994), therefore this approach does not yield a conclusive M_2 . A more reliable WD radial velocity profile may be obtained by constructing Doppler tomograms for the emission lines.

MARGARETHA PRETORIUS: Is there a radio detection of CAL 83?

ALIDA ODENDAAL: No, but an upper limit of < 0.12 mJy at 3.5 and 6.3 cm was determined by Fender, Southwell and Tzioumis (1998) from data obtained with the Australia Telescope Compact Array.