

A Study of the X-ray Variability and Accretion Mechanism in High Mass X-ray Binary Pulsars

A Thesis

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Doctor of Philosophy
in the Faculty of Science

by

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Declaration

I, Sanhita Kabiraj, hereby declare that the work presented in this thesis titled “A Study of the X-ray Variability and Accretion Mechanism in High Mass X-ray Binary Pulsars” is completely authentic. This work has been carried out by me under the supervision of Prof. Biswajit Paul (Raman Research Institute) at the Department of Astronomy and Astrophysics, Raman Research Institute under the Joint Astronomy Program hosted by the Department of Physics, Indian Institute of Science.

I further declare that the subject matter presented in this thesis has not previously formed the basis for the award of any degree, diploma, membership, associateship or any other similar title of any university or institution.

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For my dearest parents
Monoranjana Kabiraj & Sila Kabiraj

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Synopsis

High mass X-ray binaries are among the brightest X-ray sources. They host an accreting compact object (more often a neutron star) and a massive ($M > 10M_{\odot}$) donor star. They are among the earliest discovered X-ray sources and are important for various astrophysical contexts ranging from high energy astrophysics, stellar evolution, and knowledge of gravitational wave progenitors. They show variation in the X-ray intensity on all timescales ranging from seconds to days or months, which carry information about the underlying accretion mechanisms and their immediate environment. In this thesis, we have examined the observed X-ray variability in some of such sources and provided the possible physical explanations.

A brief summary of the chapters in this thesis is given below.

Chapter 1

In this Chapter, we briefly provided a background of X-ray astronomy and X-ray emission processes. We gave an overview of the classes of X-ray binary stars, accretion processes, high mass X-ray binaries and highly magnetized neutron star to provide the context of our research.

Chapter 2

Here, we described some X-ray observatories and instruments. The X-ray data from them was used to carry out the research work in this thesis. We also discussed various data analysis techniques that we have used.

In the next three major chapters, we have presented different variability studies of HMXB pulsars.

Chapter 3

In this chapter, we have investigated a unique partial eclipse in HMXB pulsar IGR J16393-4643. The eclipse profile in IGR J16393-4643 is narrow and partial, unlike other eclipsing HMXBs. The partial eclipse of IGR J16393-4643 makes it a unique candidate in the eclipsing HMXB family. We analyzed 73 *Swift*-XRT observations of this HMXB. These observations were almost evenly spread in orbital phase, providing us with complete coverage of the entire orbital phase. We extracted the lightcurves and orbital phase-resolved spectra. The soft X-ray observations also show this low-intensity phase, during which the source has about 30% of the intensity compared to the rest of the orbit. We compared the change in the spectral parameters inside and outside of the low-intensity state. Our results indicate that the partial eclipse in IGR J16393-4643 might not be an eclipse, as reported in the literature. It is possibly an outcome of the absorption in the stellar corona of the compact star grazing the neutron star. Lastly, we provided the constraints on the inclination angle of the binary for a grazing eclipse.

Chapter 4

In this chapter, we have reported our study on the Be X-ray binary GRO J2058+42. This HMXB recently went through a Type-II outburst during March-April 2019, lasting for about 50 days. It was observed by several all-sky monitors like *Fermi*, *Swift*, *MAXI*, and pointing observations were made by *NuSTAR*. With this unique opportunity, we have analyzed the broadband characteristics of the pulsar using three *NuSTAR* observations during the outburst and quiescence for the first time. We also looked at the accretion torque characteristics of the pulsar over a range of X-ray luminosity and estimated the magnetic field strength. Then, we have discussed the inadequacy of the torque-luminosity relation for the determination of the magnetic field strength of neutron stars.

Chapter 5

Here, we investigated the spin evolution of accreting persistent X-ray pulsars. The process of transfer of the accretion torque in the wind-fed persistent pulsars is complex and does not have a clear correlation with the mass accretion rate. A possible explanation of such torque behavior can be the change in the accretion mode and geometry. In this work, we have used the pulse profiles of four wind-fed pulsars as a tool to observe any changes in the accretion geometry at different accretion torque.

Chapter 6

We concluded with Chapter 6, by summarizing the implications of our results in the preceding Chapters and discussed the future prospects related to this thesis.

Publications from the Thesis

Refereed Journals

1. *An intriguing partial eclipse in HMXB pulsar IGR J16393-4643*,
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1. *A study of the pulse shape of persistent accreting pulsars at different accretion torque*
2. *Quiescent X-ray variability of GRO J2058+42*

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Chapter 1

Introduction

“Equipped with his five senses, man explores the universe around him and calls the adventure Science.”

- Edwin Hubble

1.1 First light

On 8th November 1895, the German physicist Wilhelm Conrad Röntgen discovered a previously unknown invisible ray and called it “X strahlen”, which means “X-rays”. X-rays are electromagnetic waves with a wavelength of 0.01-10 nanometers (frequency: 3×10^{16} - 3×10^{19} Hz , energy: 0.01-100 keV). At that time, the Rutherford-Bohr model of the atom was yet to be proposed which happened almost two decades later. The astronomers took almost 50 years to detect the first X-ray from an astrophysical source, the Sun ([Friedman et al. 1951](#), [Tousey et al. 1951](#)). The high energy X-ray photons interact with all the materials in their path. They get absorbed before penetrating the Earth’s atmosphere. However, the Earth is bathing with X-rays every second. Over the years, balloon and rocket experiments and satellites missions have observed several hundred thousands of X-ray sources. Hence, the “X-ray astronomy” began its journey.

1.2 X-ray Astronomy

As cosmic X-rays get absorbed by the Earth’s atmosphere, they have to be detected at the high altitudes or outer space with X-ray instruments on balloons, sounding rockets, or satellites. Therefore, X-ray astronomy emerged with the space age and is a relatively modern field as compared to optical or radio astronomy. In September 1949, the first

extraterrestrial X-ray emission from solar corona was detected by small Geiger counters aboard German V2 rocket, an experiment led by the Naval Research Laboratory ([Friedman et al. 1951](#), [Tousey et al. 1951](#)). The observed signal was weak, about one millionth of the total energy radiated by Sun in the entire electromagnetic energy band. This made the astronomers skeptical about observing X-rays from outside the solar system until the launch of an Aerobee rocket in 1962. The team led by Riccardo Giacconi was looking for the X-rays from the Moon and accidentally discovered the first X-ray Galactic source Sco X-1, located in the Scorpius constellation. The success of Aerobee started an era of X-ray astronomy. The major advancement happened in 1970 with the launch of the first X-ray satellite by NASA, named UHURU (a.k.a. Small Astronomy Satellite, SAS-A). It was the first comprehensive all sky survey in the energy band of 2-20 keV ([Giacconi et al. 1971b](#)). This mission became an outstanding scientific success with the discovery of 339 X-ray sources ([1.1](#)). Many of these sources were accretion-powered binary systems including the first three binary pulsars Cen X-3, Her X-1 and Vela X-1 ([Forman et al. 1978](#)), and also Cyg X-1, the first known black hole binary. The next milestone was set by HEAO-2 (a.k.a. Einstein) observatory that carried the first fully imaging X-ray telescope in space ([Giacconi et al. 1979](#)) and observed the sources with an angular resolution of a few arc-sec. Then several other missions like EXOSAT, Ginga and ROSAT followed by with the improvement of spectral, angular and timing resolution, field of view, energy band, and effective area.

The first X-ray mission with CCD detectors was launched in 1993, named ASCA (The Advanced Satellite for Cosmology and Astrophysics) posing a significant transformation in X-ray spectroscopy. In mid 1990s, Rossi X-ray Timing Explorer (RXTE) with excellent timing capabilities and BeppoSAX with a spectral coverage of 0.1-300 keV were launched. The subsequent X-ray missions like Chandra, XMM-Newton, Suzaku, INTEGRAL, Swift, NuSTAR, AstroSAT, NICER have revolutionized the understanding of today's X-ray sky. The Extended Roentgen Survey with an Imaging Telescope Array (eROSITA) and Imaging X-ray Polarimetry Explorer (IXPE) have been recently launched. At present several hundred thousand X-ray sources are known and at least several hundred individual sources have been studied extensively. Upcoming missions like POLIX, XRISM, eXTP, ATHENA etc. will carry out polarimetry and high-resolution spectroscopy. In the last 60 years since the discovery of Sco X-1, X-ray astronomy has matured over time and has yet to go a long way.

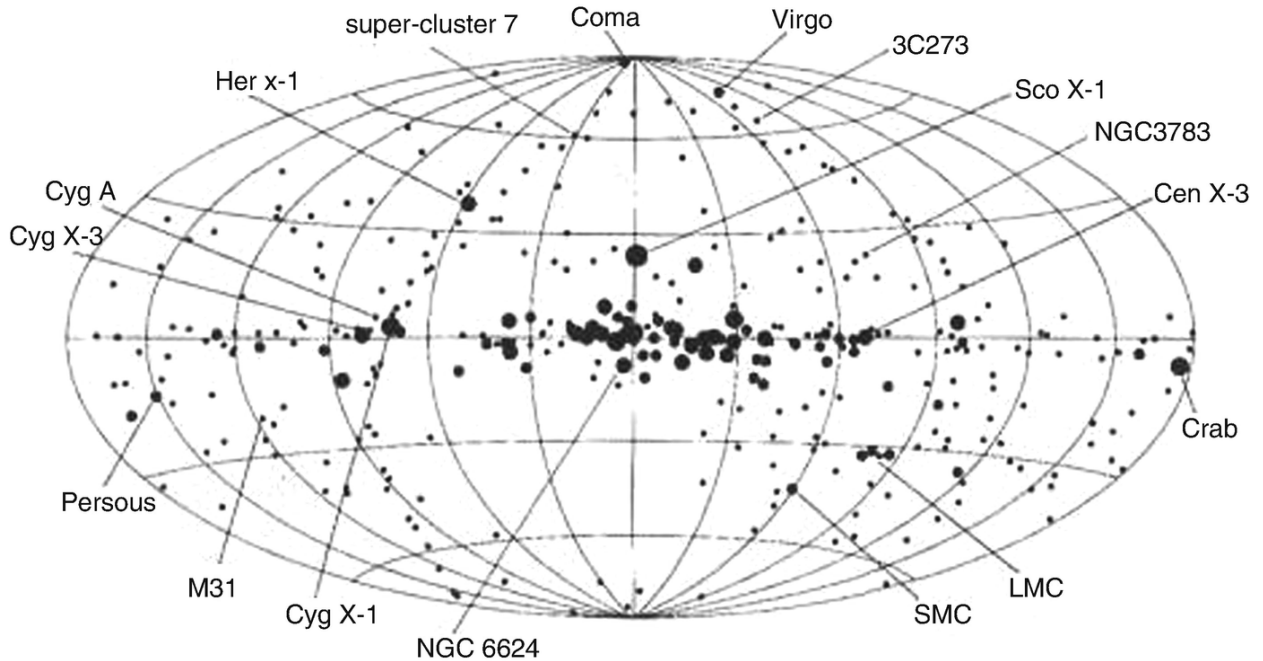


Figure 1.1: The observed X-ray sources by Uhuru are represented by dots in galactic coordinates. The size of the dots are proportional to their intensity on a logarithmic scale. Image credit: X-ray Astronomy (Eds. R. Giacconi, H. Gursky), [1974](#), Riedel, Dordrecht, p. 156.

1.3 The X-ray emission processes

The X-rays from the celestial sources may originate from thermal or non-thermal emission processes and also from atomic electronic transitions.

1.3.1 Thermal emission

Black-body radiation : It occurs when a body is in thermodynamic equilibrium with its surrounding. The black-body spectrum is a Planck distribution. The temperature of the body determines the peak of the spectrum of the emitted radiation. Anything that has a temperature around $10^7 K$ can emit thermal X-ray emission.

The emission from optically thick accretion disk of the X-ray binaries is multi-temperature black-body radiation originating from the different radius of the accretion disk.

Bremsstrahlung radiation : When acceleration of charged particles (often electrons) is caused by electric field of other charged particles (often protons or atomic nucleus), it produces Bremsstrahlung radiation or “free-free” radiation. If the energy of the incident charged particles or electrons is enough, X-ray photons are emitted. The spectrum of this

radiation is continuous.

All optically thin hot plasma produces X-ray bremsstrahlung emission. This radiation can also be detected as diffuse X-ray emission from hot gas in the galaxy clusters with temperatures of the order of 10 million Kelvin.

1.3.2 Non-thermal emission

Synchrotron radiation : This occurs when charged particles are accelerated along a spiral path under the influence of a magnetic field. The magnetic field exerts a force that is perpendicular to the field and the direction of the velocity of the particles. This causes the particles to follow a spiral path around the magnetic field lines. High energy photons are produced when the electrons have sufficiently high energy, or the magnetic field has sufficiently high strength. The emission spectrum is the combined result of the the emission spectra of individual electrons which follows a power-law decay. Another important characteristics of this radiation is polarization.

Synchrotron radiation at radio wavelengths is seen from the supernovae remnants, jets and many other sources with strong magnetic field. The optical and X-ray emission from Crab pulsar is known to originate from synchrotron radiation.

Inverse Compton scattering : When low energy photons are scattered by high energy electrons and gains energy, that process is known as inverse compton scattering. It is 'inverse compton' because the electrons lose energy instead of gaining. The spectra produced by this process has power law shape.

It is a key process of X-ray emission from active galaxies and black holes. The power law spectra of high mass X-ray binaries can be result of inverse compton scattering of soft X-ray photons by high energy electrons in the vicinity of the compact object ([Hayakawa 1985](#)).

1.3.3 Spectral lines

The spectral lines are often present along with the continuum of the X-ray spectra of binaries. They may occur due to fluorescence or recombination.

Fluorescence lines : When an atom is excited by collision with X-ray photon, one of the electrons can be removed from the inner most orbit creating a hole. Electron from an

outer orbit falls into the hole and the difference of the binding energies of the two orbits is emitted as fluorescent X-ray. This emitted energy is a characteristic of the element. The iron K-shell lines (Fe $K\alpha$ and Fe $K\beta$ lines) are often present in the spectra of X-ray sources. They provide an insight to the density and ionization of the matter surrounding the compact object. Sometimes asymmetric broadening of these lines are seen as a result of the gravitational redshift. The iron fluorescence lines were first detected in several AGNs and may provide a measure of the spin of the black holes in binaries and AGNs. Fluorescent X-ray emissions from various other elements like Si, Mg, O, Ne etc. have also been detected.

Charge exchange emission : When a charged ion collides with a neutral atom, the atom becomes unstable by transferring one or more electrons to the ion. Electrons from the outer shell of the atom falls into a lower shell to make it stable and photons are emitted. Highly charged C, O and Ne can produce X-ray emission by this process. Charge exchange of the ions in the solar wind with neutral atoms in the interstellar medium and the Earth's geocorona contributes to the soft X-ray background.

1.4 Accretion

Accretion is the most efficient and powerful process that causes high energy radiation from the binary X-ray sources. It is the process of accumulation of gas or matter by a celestial body due to its gravitational attraction. If a spherically symmetric body of mass M and radius R accretes a mass m ($m \ll M$), the exerted gravitational potential energy will be,

$$\Delta E_{GR} = \frac{GMm}{R} \quad (1.1)$$

If the accumulation rate is $\dot{M}$ and all of this energy is radiated, then the luminosity of the body will be,

$$L_{acc} = \frac{GM\dot{M}}{R} \quad (1.2)$$

where G is the gravitational constant. The quantity $\dot{M}$ is called the mass accretion rate and L_{acc} is called the accretion luminosity. The efficiency of accretion is strongly dependent upon the factor $\frac{M}{R}$, i.e., the compactness of the object. The accretion efficiency is expressed as,

$$\eta = \frac{GM}{Rc^2} \sim \frac{R_s}{2R} \quad (1.3)$$

where c is the speed of light in vacuum and $R_s = \frac{2GM}{c^2}$ is the Schwarzschild radius. For a typical neutron star with mass $\sim M_\odot$ and radius $\sim 10\text{km}$, η is ~ 0.15 . The observed luminosity from an accreting neutron star is roughly $10^{36} - 10^{38} \text{erg s}^{-1}$. It requires a mass accretion rate of $\sim 10^{-10} - 10^{-8} M_\odot \text{yr}^{-1}$. The increase in the mass accretion rate causes an increase in the luminosity. At a certain luminosity the outward radiation pressure on the electrons balances the inward gravitational force on the protons. It is the maximum luminosity that can be achieved by a body and known as the Eddington luminosity. It is given by,

$$L_{Edd} = \frac{4\pi GMm_p c}{\sigma_T} \simeq 1.3 \times 10^{38} \frac{M}{M_\odot} \text{erg s}^{-1} \quad (1.4)$$

where m_p is the mass of proton and σ_T is the Thomson scattering cross-section. For spherically symmetric accretion, the Eddington accretion rate will be,

$$\dot{M}_{Edd} = \frac{L_{Edd}}{\eta c^2} \quad (1.5)$$

When $L > L_{Edd}$, the radiation pressure halts further accretion. However, there are transient events like supernova or X-ray bursts when we observe luminosity that exceeds L_{Edd} .

Accretion is a common phenomena in the universe and hence an important process in astrophysical context. Stars sweeping up interstellar dust or Earth being bombarded by meteorites are all examples of accretion. For accretion to produce high energy emission a compact accreting star has to be present. X-ray binaries and super-massive black holes at the center galaxies are such systems.

1.5 Classification of X-ray binaries

X-ray binary systems host at least one degenerate star. It can be a black hole, a neutron star or a white dwarf that rotates with its optical companion centering the common center of mass. Based on the nature of the this degenerate star or the compact object, the X-ray binaries can be categorized in three groups, 1) Black hole X-ray binary (BH-XRB), 2) Neutron star X-ray binaries (NS-XRB), and 3) White dwarf X-ray binary (WD-XRB) or Cataclysmic Variable (CV).

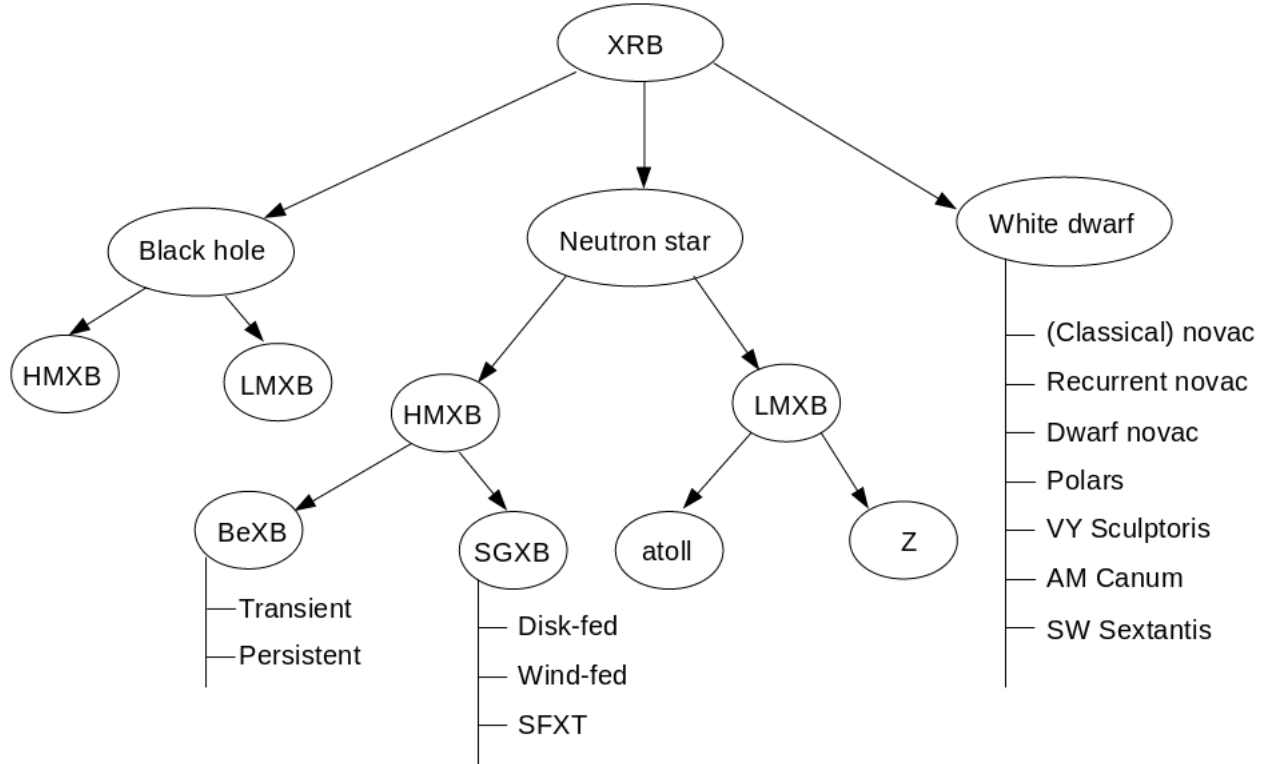


Figure 1.2: The classification of the X-ray binaries (Reig 2011)

The commonly used classification of XRBs is based on the mass of the companion. If the mass of the companion is $\geq 10M_{\odot}$, it is an High Mass X-ray Binary (HMXB) and if the companion star has a mass $\leq 1M_{\odot}$, then it is a Low Mass X-ray Binary (LMXB). This phenomenological classification is restricted to the NS and BH-XRBs. X-ray binaries with a companion mass between the above mentioned classes are extremely rare and are called the Intermediate X-ray binaries.

1.5.1 HMXB

The optical companion of HMXBs are usually bright massive OB stars. These stars strongly radiate in the optical band and the optical luminosity (L_{opt}) is higher than the X-ray luminosity (L_x). The average age of HMXBs depends on the evolutionary time scale of the OB star, which typically lasts for $10^5 - 10^7$ yr. HMXBs are young galactic population and are distributed in the star forming region of the galaxy. In the Milky Way, they are concentrated around the galactic plane and the spiral arms. Such systems have been observed from other galaxies as well.

HMXBs often host neutron star as the compact object. However, there are few known to host black holes. The neutron star in a HMXB is young and has a high magnetic field ($\sim 10^{12}$ Gauss), which governs the flow of the accreted matter. The matter gets channeled through the magnetic field lines, falls onto the poles of the neutron star and then X-ray radiation is emitted. Therefore, such HMXBs are observed as pulsars and they are the focus of this thesis.

Based on the type of the companion star HMXBs can be divided into two groups : **Supergiant X-ray Binary (SGXRB)** with a OB supergiant, and **Be X-ray Binary (BeXRB)** with a Be star (B-type emission line star). SGXRBs have shorter orbital period (usually less than 10 days) and smaller orbital eccentricity. The orbital period of BeXRBs is longer which can vary from months to years and the orbit is highly eccentric. The supergiant star of an SGXRB is either in the *Roche Lobe* filling stage or has already filled the *Roche Lobe*. The mass transfer from the supergiant happens accordingly, either via stellar wind or via *Roche Lobe* overflow. These mass transfer processes have been described in the later section (1.6).

1.5.2 LMXB

The companion in an LMXBs is a late spectral type (K, M) main-sequence star (or a white dwarf). LMXBs emits most of its radiation in X-rays and very little (less than one percent) in visible light. They are among the old stellar population and distributed in the globular clusters and the Galactic bulge. Their average life span is longer ($\sim 10^7 - 10^9$ yr) and the orbital period ranges from minutes to days. The companion does not have a strong stellar wind and the mass is transferred via *Roche Lobe* overflow.

The compact star can be a black hole or a neutron star. The neutron star in a LMXB has weaker magnetic field ($\sim 10^9 - 10^{11}$ Gauss), as the magnetic field decays with the evolution of the neutron star. The weak magnetic field can not channel the accreted matter along the field lines. As a consequence, the neutron stars in LMXBs rarely show X-ray pulsations.

1.6 Accretion mechanisms

The mass transfer on to the compact star in an X-ray binary can happen by one of the three following mechanisms.

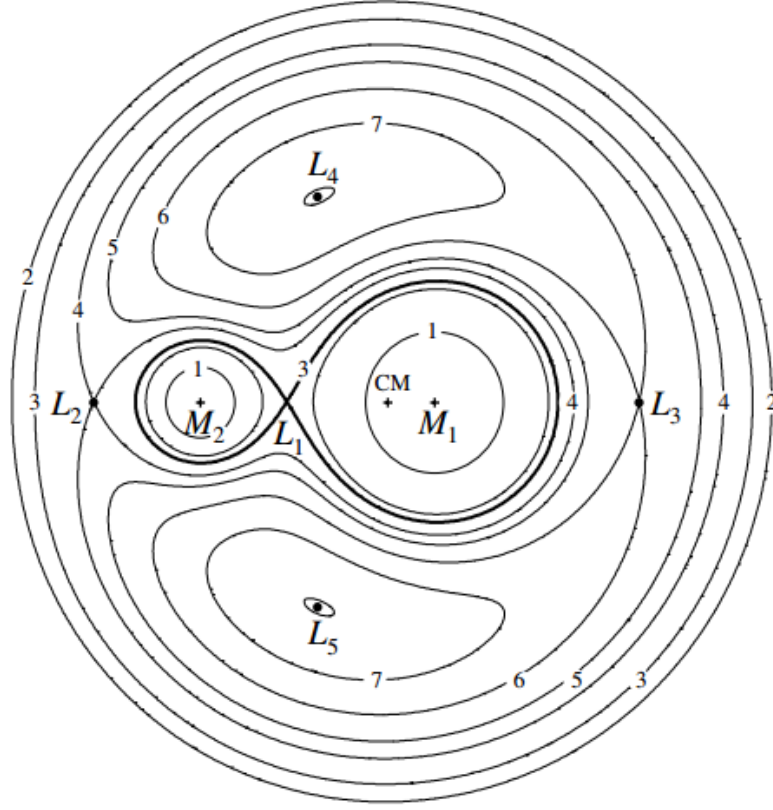


Figure 1.3: The orbital plane of the *Roche* equipotentials ($\Phi_R = \text{constant}$) for a binary system with mass ratio $q = 0.25$ (Frank et al. 2002).

1.6.1 Roche Lobe overflow

To understand this, let us consider the motion of a test particle of mass m in the gravitational potential of two stars with masses M_1 and M_2 revolving in a circular orbit around their center of mass. In the co-rotating frame with the binary, the mass m is subjected to the gravitational potential of both the stars and the centrifugal force. If the position vector of the center of the stars are $\mathbf{r}_1$ and $\mathbf{r}_2$, then the *Roche potential* is,

$$\Phi_R(\mathbf{r}) = -\frac{GM_1}{|\mathbf{r} - \mathbf{r}_1|} - \frac{GM_2}{|\mathbf{r} - \mathbf{r}_2|} - \frac{1}{2}(\boldsymbol{\omega} \times \mathbf{r})^2 \quad (1.6)$$

where $\boldsymbol{\omega}$ is the orbital angular velocity and $\mathbf{r}$ is the position vector of the test mass. As we are considering the rest frame of the binary, the *Roche potential* Φ_R does not include the Coriolis force. Figure 1.3 shows the equipotential surfaces of Φ_R for $q = \frac{M_1}{M_2} = 0.25$. The mass ratio q determines the shape of the equipotentials and the binary separation

a determines the scale. The labels 1-7 of the equipotentials are given in the increasing order of Φ_R . L_1 , L_2 , L_3 , L_4 and L_5 are the five ‘Lagrangian’ points where the gradient of the effective potential is zero. The regions surrounding both the stars that create a figure-of-eight area (the bold line) are known as *Roche Lobes*. In three dimension this figure-of-eight area is actually a dumbbell like structure and the two *Roche Lobes* are connected at the inner Lagrangian point L_1 . *Roche Lobe* of a star is the region where matter is gravitationally bound to that star. The star can expand only until its surface is within this region. If one of the two stars in the binary fills its *Roche Lobe* and expands further then the matter will flow to the other star through L_1 and will be captured by it. This mass transfer process is called the *Roche Lobe overflow*.

The matter that passes through L_1 has high specific angular momentum which stops it from directly falling onto the surface of the second star. The infalling matter moves in circular orbits around the second star and spirals in towards it forming an Accretion disk. Accretion via disk is a dominant mechanism of mass transfer in the LMXBs and also in some of the HMXBs. When the accreting star is neutron star the dynamics of the accretion depends mainly on the spin and magnetic field of the neutron star.

1.6.2 Stellar wind accretion

This type of accretion is witnessed in the binary systems with a close orbit containing a massive highly luminous early-type (O or B) star and a neutron star (or a black hole). The majority of wind accretors are actually accretion powered pulsars.

The early-type star (donor star) loses mass in the form of intense and super-sonic stellar wind driven by its radiation pressure. For steady homogeneous spherically symmetric wind, the mass loss rate is,

$$\dot{M}_w = 4\pi r^2 \rho(r) v_w \quad (1.7)$$

where ρ is the density of the ejected matter at a distance r . v_w is the wind velocity, which is about a few thousands $km\ s^{-1}$ and is much greater than the sound speed. Only a fraction of this stellar wind is captured by the neutron star. When the matter in the wind comes close to the compact object and the gravitational potential energy exceeds the kinetic energy of that matter, then it gets accreted. The accretion happens in a roughly cylindrical region of radius r_{acc} with its axis along the relative direction of the wind. r_{acc} is called the accretion radius and expressed as,

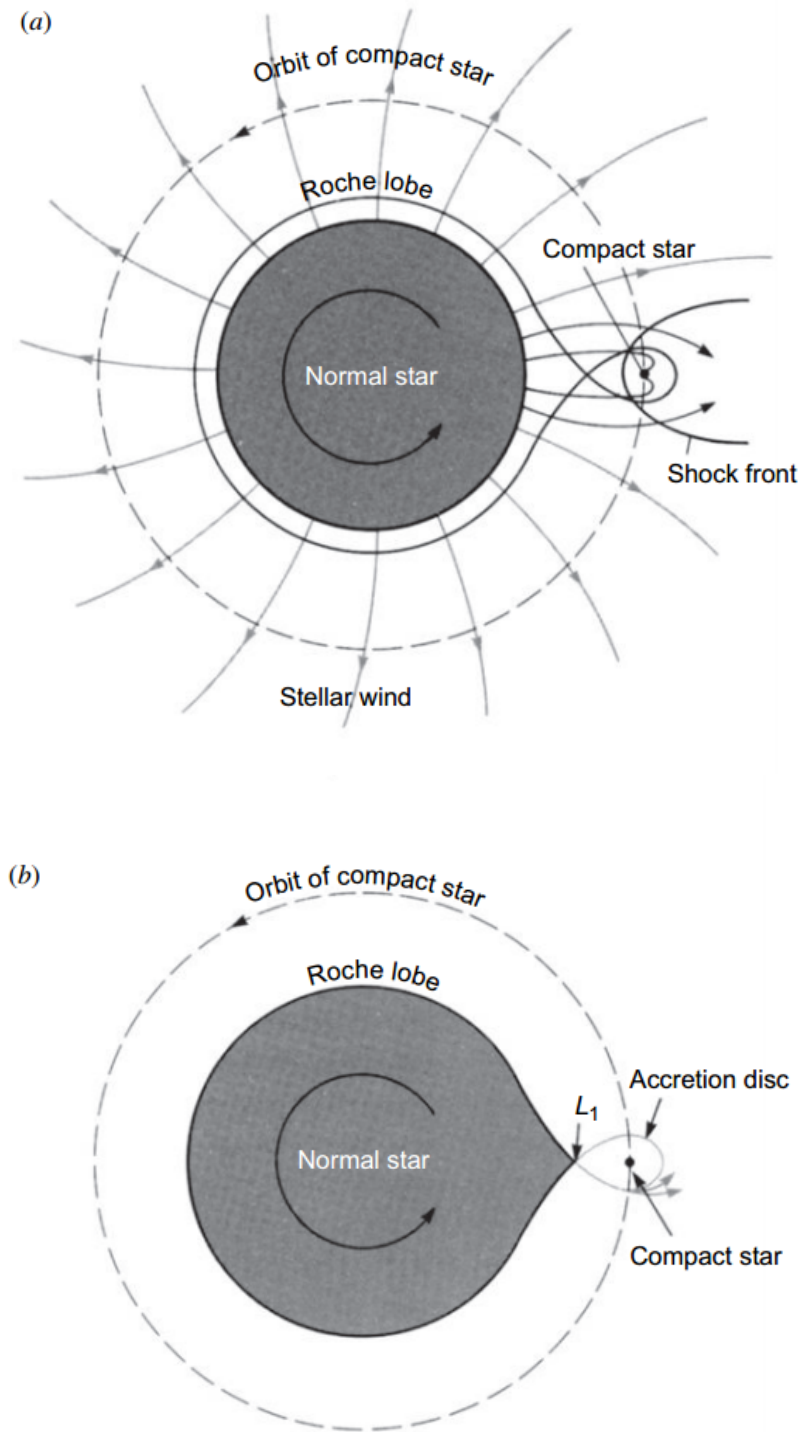


Figure 1.4: A rough sketch of two accretion processes in X-ray binaries. In (a), the compact object is embedded in the stellar wind of the massive companion and accreting from it. In (b), the matter from the companion is being transferred through the Lagrangian point L_1 (Longair 2011).

$$r_{acc} \simeq \frac{2GM_n}{v_{rel}^2} \quad (1.8)$$

where, $v_{rel} = (v_n^2 + v_w^2)^{1/2}$ is the relative wind velocity. M_n and v_n are the mass and the orbital velocity of the neutron star (typically $\sim 200 \text{ km s}^{-1}$). As the wind speed (v_w) is highly supersonic, a bow shock is formed near the neutron star. The mass accretion rate and the angular momentum carried by the wind material are the deciding factors for the formation of an accretion disk around the neutron star.

Assuming, $v_w \gg v_n$, v_{rel} is almost equal to v_w . Then the mass fraction η i.e. accreted by the neutron star can be expressed in terms of the mass accretion rate $\dot{M}_{acc}$ and the mass loss rate of the companion as,

$$\eta = \frac{\dot{M}_n}{\dot{M}_w} \simeq \frac{\pi r_{acc}^2}{4\pi a^2} = \frac{G^2 M_n^2}{a^2 v_w^4} \quad (1.9)$$

Considering the wind velocity has to be at least around the escape velocity in order to eject mass, $v_w \sim v_{esc} = \sqrt{2GM_c/R_c}$ (where M_c and R_c are the mass and radius of the companion). Then equation 1.9 becomes,

$$\eta = \frac{1}{4} \left(\frac{M_n}{M_c} \right)^2 \left(\frac{R_c}{a} \right)^2 \quad (1.10)$$

Equation 1.10 implies that the accretion from stellar wind is an inefficient process ($\eta \sim 10^{-4}$). In such systems, the stellar wind is intense and the mass loss rate is typically $10^{-6} - 10^{-5} M_\odot \text{ yr}^{-1}$ ($\sim 10^{20} - 10^{21} \text{ g s}^{-1}$). Therefore, the observed luminosity in the wind accreting binaries lies between $10^{35} - 10^{37} \text{ erg s}^{-1}$.

1.6.3 Accretion from Be star

BeXRBs constitute about two third of the identified HMXBs. The Be optical companion has the luminosity class of III-V, same as the B stars with strong emission of hydrogen (Balmer emission lines). They produce strong infrared radiation, which is known as infrared excess. This infrared excess and the emission lines originate from the equatorial circumstellar disk of ionized material around the Be star. The disk is formed by the ejected material by the star and the formation mechanism remains an open question.

The Be star stays well inside its *Roche Lobe* and the binary orbit is wide and eccentric. So, for most of the orbital phase the compact star remains at a large distance from its companion. The mass ejection from rapidly rotating Be star occurs erratically. When the

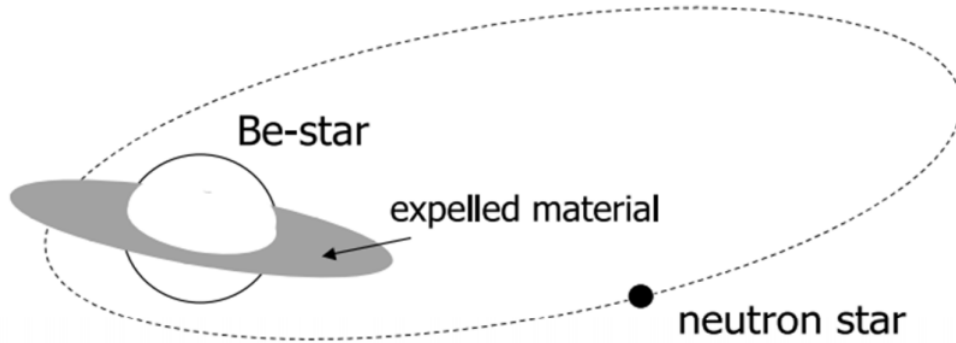


Figure 1.5: A schematic diagram of a Be X-ray binary with a neutron star. The ejected matter from the Be star forms a circumstellar disk and the neutron star accretes matter from it.

neutron star passes through the circumstellar disk of the Be star or crosses the periastron passage, accretion sets in. The capture of matter from the companion happens at the accretion radius r_{acc} and gets accreted in a similar fashion as in the case of wind-fed HMXBs. The abrupt accretion results in a sharp increase in the X-ray emission, which is known as outburst. Therefore, BeXRBs remain most of the time in the quiet state and then go through bright transient X-ray outbursts.

There are two types of outbursts seen in BeXRB, Type I and Type II (giant) outburst. Type I outbursts are periodic, separated by the orbital period. They peak when the neutron star is close to the periastron passage. The luminosity during these type of outbursts is usually $10^{36} - 10^{37} \text{ erg s}^{-1}$. Type II or giant outbursts are more luminous ($L_x \geq 10^{38} \text{ erg s}^{-1}$). The luminosity can even exceed the Eddington luminosity of the neutron star (Vasilopoulos et al. 2020). These outbursts last for a large fraction of the orbital phase or for several orbital periods and do not show orbital phase dependence. In chapter 4, we have reported results from observations during one of such giant outburst from a BeXRB GRO J2058+42.

1.7 Evolution of HMXBs

The basic model of formation and evolution of HMXBs was developed over the years (Tutukov and Yungelson 1973, van den Heuvel and De Loore 1973, Flannery and van den Heuvel 1975, van den Heuvel 1976, Postnov and Yungelson 2014, van den Heuvel

2019).

The formation process starts with two main sequence star with high initial mass. Figure 1.6 describes this evolution process with two main sequence star with mass $M_1 = 14.4M_\odot$ and $M_2 = 8M_\odot$ and an orbital period of ~ 100 days.

1. It starts with the zero-age main sequence (ZAMS). At this stage, a pair of detached high mass main sequence OB stars are well inside their Roche Lobe. They evolve for about few million years when the tidal interaction is very effective. The time of core hydrogen burning is $t_{nuc} \sim M^{-2}$. The more massive star M_1 burns out its hydrogen and forms a dense helium core as it leaves the main sequence. It starts to expand rapidly and eventually fills its Roche lobe.

2. As the more massive star fills its Roche lobe, mass is transferred to the less massive star M_2 which is still in main sequence. At this point, if the mass difference between the two stars is small then the mass transfer will happen via a non-conservative yet stable Roche lobe overflow. Due to the mass transfer, M_2 can accumulate large angular momentum.

3. The mass transfer is done when the entire hydrogen envelope of M_1 is lost and only a helium core is left. This naked core can be observed as a Wolf Rayet (WR) star. A close binary forms in the process of mass transfer and M_2 is now more massive. Both the stars evolve independently and eventually M_1 goes through supernova explosion.

4. As a result of the supernova explosion, M_1 becomes a neutron star (or a low mass black hole). The physical conditions that decide the formation of a neutron star or a black hole are still unknown. The explosion does not disrupt the system if less than half of the total mass of the system is ejected (Blaauw 1961). The bound system achieves a longer orbital period and an altered eccentricity.

5. After evolution of the binary for about few million years, the massive companion (M_2) gradually starts to fill its Roche lobe. The compact object accretes matter from its massive companion in the form of stellar wind or Roche lobe overflow. Thus the binary evolves into a HMXB system.

6. If the compact object is a neutron star, its evolution depends on the spin and the surface magnetic field strength. The massive companion eventually fills its Roche lobe. All the HMXBs with neutron star enters into the common envelope evolution (CEE) phase as an outcome of unstable Roche lobe overflow. These systems survive the CEE only if the orbital period at the start of this stage is longer than about one year (Taam 1996). In that case, the CEE stage ends with a formation of a helium star or WR star and a neutron star. Otherwise, the neutron star merges with the helium core during the common envelope

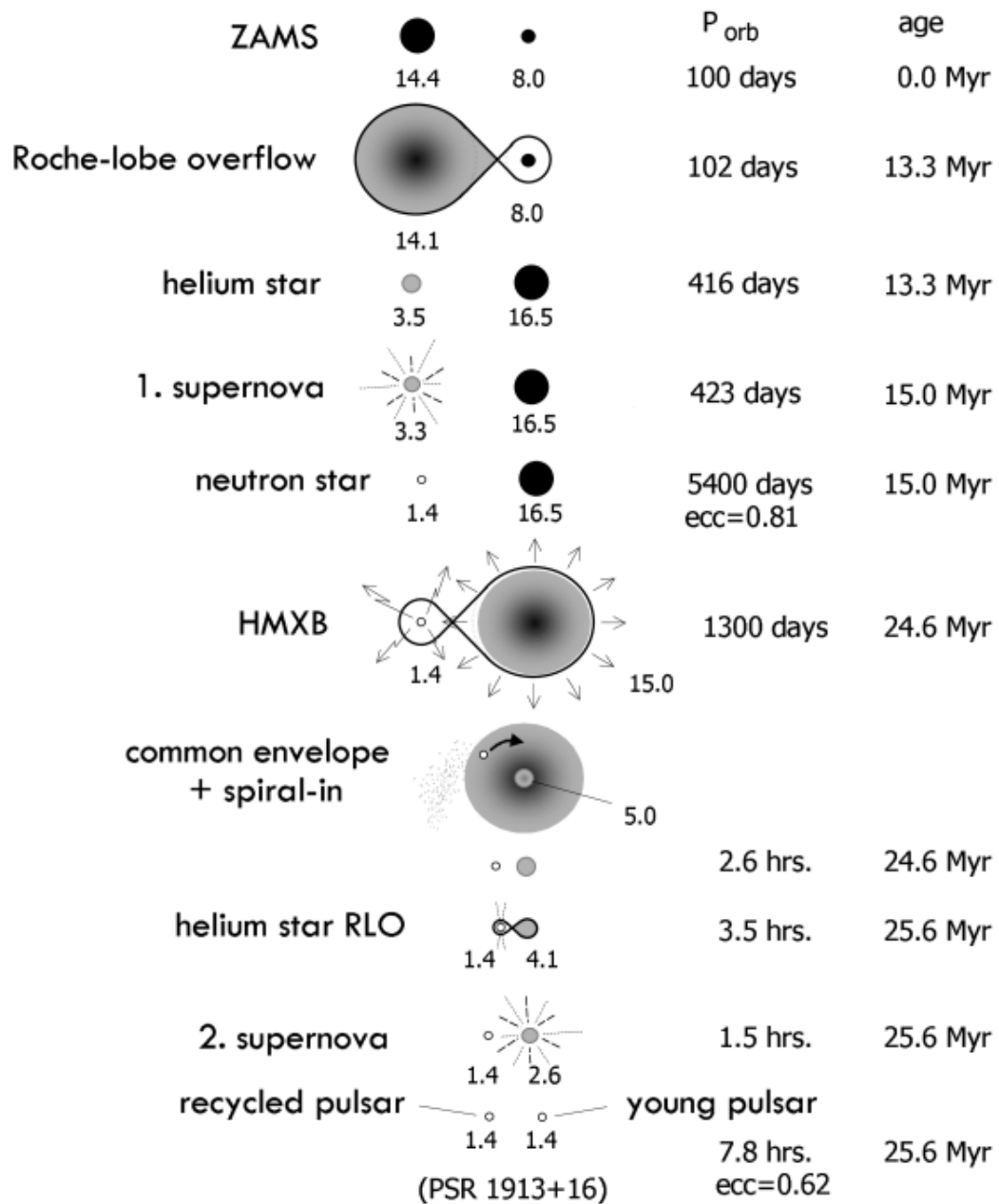


Figure 1.6: The formation and evolution of neutron star HMXB ([Tauris and Van Den Heuvel 2010](#)).

and are expected to form the hypothetical Thorne-Zytkow star (Thorne and Zytkow 1977).

7. If the system survive with a helium star and a neutron star, then after the supernova explosion of the helium star it is possible to form a close double neutron star system.

Here we have mainly discussed the evolution of a HMXB with a neutron star as this thesis is focused on them.

1.8 Accretion onto highly magnetized neutron stars

X-ray binary pulsars are binary systems with a highly magnetized neutron star ($\sim 10^{12}$ Gauss). The study of accretion powered pulsars began with the discovery of Cen X-3 in 1971 (Giacconi et al. 1971a). It showed a clear pulse period of ~ 4.8 seconds, which was longer than that of the radio pulsars. More than 160 accreting pulsars have been discovered since then. The pulsations are attributed to the rotating beam of radiation from the poles of the neutron star due to the misalignment of the magnetic dipole axis and the spin axis. With the assumption of a dipole geometry of the magnetic field of the neutron star, the magnetic field strength at a distance r will be,

$$B \sim \frac{\mu}{r^3} \quad (1.11)$$

where μ is the magnetic moment of the neutron star. Therefore, magnetic pressure exerted on the accreted matter is,

$$P_{mag} \sim \frac{\mu^2}{8\pi r^6} \quad (1.12)$$

Considering spherically symmetric accretion, the magnetic pressure will balance the ram pressure of the infalling matter at a certain radius, which is called the Alfven radius (or, the magnetospheric radius). At Alfven radius r_A ,

$$\frac{\mu^2}{8\pi r_A^6} = \rho v^2 = \frac{(2GM)^{1/2} \dot{M}}{4\pi r_A^{5/2}} \quad (1.13)$$

Here ρ and v are the density and velocity of the infalling matter. v is comparable to the free fall velocity, $\sqrt{\frac{2GM}{r}}$. M is the mass of the neutron star and $\dot{M}$ ($= 4\pi\rho v^2$) is the mass accretion rate. Therefore from equation 1.13 we get,

$$r_A = 2.9 \times 10^8 M^{1/7} R_6^{-2/7} \mu_{30}^{4/7} L_{37}^{-2/7} \text{ cm} \quad (1.14)$$

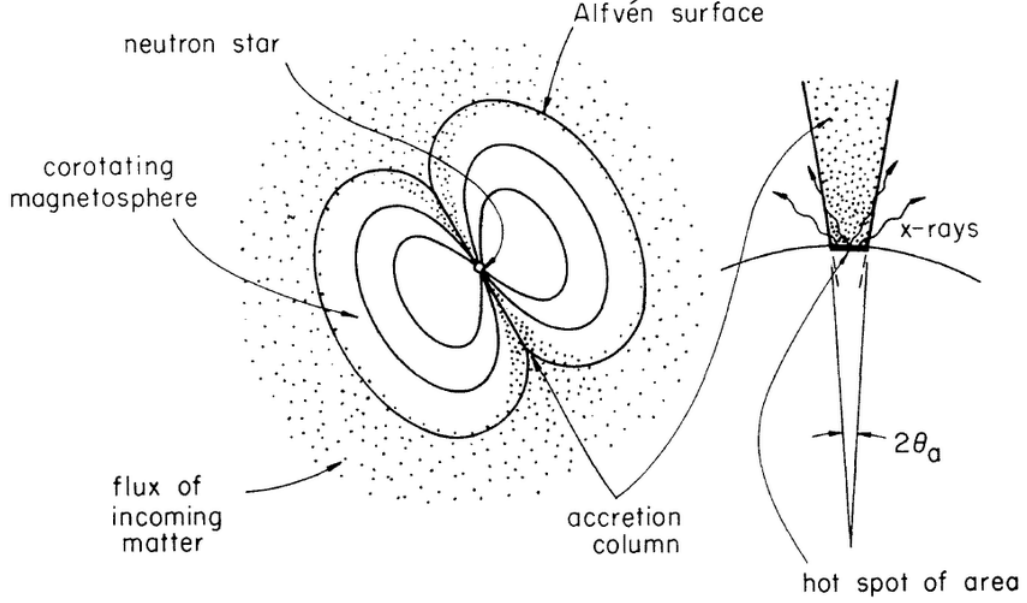


Figure 1.7: A rough sketch of accretion on a highly magnetized neutron star. The left side shows an overall view of the magnetosphere and the flow of the accreted matter. The right side is the close-up view of the accretion column near the neutron star surface (Lamb et al. 1973).

where R_6 , μ_{30} and L_{37} are the radius, magnetic moment and the luminosity in the units of 10^6 cm, 10^{30} G cm³ and 10^{37} erg s⁻¹.

Another characteristic radius is the corotation radius. It is the radius at which the spin angular velocity (ω) of the neutron star equals the Keplerian velocity of the infalling matter. It is given by

$$r_{co} = \left(\frac{GM}{\omega^2} \right)^{1/3} = 1.5 \times 10^8 M^{1/3} P_s^{2/3} \text{ cm} \quad (1.15)$$

where P_s is the spin period.

The interplay of the three characteristic radii r_A , r_{co} and r_{acc} (equation 1.8) determines the flow of the accreted matter onto the neutron star. They are the functions of the wind velocity, mass accretion rate and spin period. These parameters vary causing the change in the relative lengths of the three radii. When $r_A > r_{acc}$, the stellar wind material surrounding the neutron star interacts with the magnetosphere and does not experience enough gravitational pull. This causes the *magnetic inhibition* to the accretion. If $r_{acc} > r_A > r_{co}$, the wind material manages to penetrate through the accretion

radius. But the centrifugal force inhibits the matter to penetrate further into the magnetosphere. The magnetic field imposes a super-Keplerian drag on the wind material. It is called the *centrifugal inhibition* or the *propeller effect*. When $r_{acc} > r_A$ and $r_{co} > r_A$, the wind material crossing the r_{acc} penetrates through the magnetosphere, gets channelled along the magnetic field lines to the magnetic poles of the neutron star. Therefore, this is known as the *accretion regime*.

The infalling material creates column shaped structure at the poles. It is called the accretion column. [Basko and Sunyaev \(1976\)](#) proposed theoretical models for different configurations of the accretion column by taking into account different hydrodynamic effects at different accretion rates. Changes in the accretion rate transforms the emission diagram or the beaming pattern. When the accretion rate is high, a radiation dominated shock is formed at a certain distance above the surface of the neutron star, which causes deceleration of the plasma to subsonic speed. The very hot plasma subsequently sinks into a cylindrical region on the surface, emitting X-rays. An accretion column raises and becomes optically thick. The X-ray photons escape through the side walls of the accretion column resulting a broad emission geometry with a *fan beam* pattern. However, at lower accretion rate, the radiative shock is not formed and infalling plasma is decelerated by Coulomb interactions and collective plasma oscillations. The X-ray photons can escape along the column axis as the optical depth is less than 1, creating a *pencil beam pattern*. Switching between these two pattern occurs at a critical luminosity L_{crit} when the radiation dominated shock becomes effective. For typical neutron star parameters, [Becker et al. \(2012\)](#) calculated

$$L_{crit} = 1.5 \times 10^{37} B_{12}^{16/15} \text{ erg s}^{-1} \quad (1.16)$$

where B_{12} is the surface magnetic field strength in the unit of $10^{12} G$. Later, [Mush-tukov et al. \(2015\)](#) provided a more detailed prescription for L_{crit} by considering Compton scattering including the resonances, polarization composition, the temperature of the infalling matter and the geometry of the flow, which is a non-monotonic function of the magnetic field.

1.8.1 Pulse profiles

Both the accretion geometry and the emitted radiation are strongly anisotropic. This emitted radiation appears pulsed due to the misalignment of the magnetic dipole axis

with the spin axis. The observed modulated intensity with the rotation of the neutron star is described as the pulse profile. The X-ray binary pulsars have complex pulse profiles with a wide variety of shapes and amplitudes. They are the unique characteristics of the pulsars. The shaping of the pulse profile of a pulsar depends on various factors like the orientation of the magnetic dipole axis and spin axis with respect to the observer's line of sight, the accretion rate, the geometry of the accretion, the beaming pattern of the emitted radiation, the magnetic field configuration and the gravitational light bending. Depending on the visibility of the hotspots (magnetic poles), the profiles can be single peaked, double peaked, or sometimes multiple peaked (Bulik et al. 2003). Several theoretical models have been developed for emission regions and beaming patterns to explain and reproduce the observed pulse profiles (Kraus et al. 2003, Leahy 1991). However, the exact morphology remains unknown.

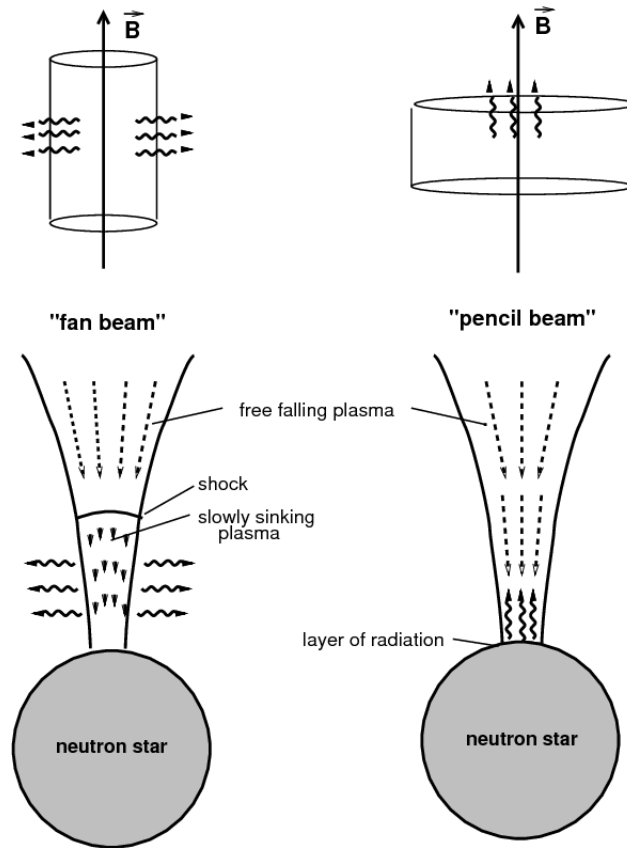


Figure 1.8: A schematic diagram of the accretion geometry and the formation of the radiation pattern. The left side shows the *fan beam* pattern at the super-critical luminosity and the right side shows the *pencil beam* pattern at the sub-critical luminosity (Schönherr et al. 2007).

The observations of the pulse profiles of the accreting pulsars often show strong dependence on the X-ray luminosity and X-ray energy. The pulse profiles can change at the critical luminosity as the beaming pattern changes from *pencil beam* to *fan beam*. In BeXRB pulsars, a drastic increase in the mass accretion rate causes a sharp rise in the X-ray luminosity as the binary goes to an outburst phase from the quiescent. It incorporates a change in the beaming pattern, and therefore in the pulse profiles. An example of this is EXO 2030+375 (Parmar et al. 1989a).

At lower energies, the X-ray photons suffer scattering and absorption in the accretion column, which results in a complex shape. Sometimes, dip-like features (X-ray dips) are seen at certain pulse phases because of the absorption in the optically thick material in the accretion column. At higher energies, the pulse profiles exhibit simpler shapes and reflect the intrinsic beaming pattern.

1.8.2 Cyclotron resonance scattering feature

X-ray binary pulsars often display broad absorption lines in the spectra. The strong surface magnetic field quantizes the plasma electrons to their Landau levels and the resonant scattering of the X-ray photons with these quantized electrons results in line like absorption features, which is called the *cyclotron resonance scattering features* or the *cyclotron lines*. Their energies provide a direct measurement of the surface magnetic field strength of the neutron star.

Let us consider, an electron with charge e is moving in a helical trajectory under the influence of magnetic field B . The radius of gyration (or the Larmor radius r_L) and the Larmor frequency ω_L are,

$$r_L = \frac{m_e v_{\perp}}{eB} \quad (1.17)$$

$$\omega_L = \frac{eB}{m_e} \quad (1.18)$$

where m_e is the mass of the electron and $v_{\perp}$ is the perpendicular velocity component. The relativistic quantum mechanical effects become important when the Larmor radius approaches the *de Broglie* wavelength, i.e.,

$$\lambda_{deBroglie} = \frac{\hbar}{m_e v} \quad (1.19)$$

where $\hbar$ is the reduced Planck's constant. In this scenario, the perpendicular kinetic energy of the electrons with respect to the direction of the magnetic field becomes quantized to their Landau levels, which is given as

$$\frac{p_{\perp}}{m_e c^2} = n \frac{B}{B_c} \quad (1.20)$$

where, c is the speed of light. $B_c (= \frac{m_e^3 c^3}{e \hbar}) \sim 4.4 \times 10^{13} \text{ G}$, is the critical value of the magnetic field when the kinetic energy of the electron approaches its rest mass. Therefore, the scattering cross-section of the photons becomes resonant at the Landau energy levels of the electrons in the plasma. The photons with the energy equivalent to the first Landau energy level, or its multiples, become missing from the line of sight, which results a line like absorption feature in the emerging X-ray spectra. The energy of the absorption feature is given by,

$$E_n = (m_e c^2) \frac{\sqrt{1 + 2n \frac{B}{B_c} \sin^2 \theta} - 1}{\sin^2 \theta} \times \frac{1}{1 + z} \quad (1.21)$$

where θ is the angle between the direction of the incident photon and the direction of the magnetic field. z is the gravitational redshift at the neutron star surface. It is approximately

$$z = \frac{1}{\sqrt{1 - \frac{2GM}{Rc^2}}} - 1 \quad (1.22)$$

In the non-relativistic approximation, the first energy level ($n = 1$) is

$$E_n = n \frac{\hbar e B}{m_e} \sim 11.6 \frac{B}{10^{12} \text{ G}} \quad (1.23)$$

Here, $n = 1, 2, 3, \dots$, represents the number of the harmonic. The line corresponding to $n = 1$ is the fundamental cyclotron line and $n \geq 2$ are the harmonic lines. Considering the redshift, the observed fundamental cyclotron line energy will be

$$E_{cyc} = E_n (1 + z)^{-1} = 11.6 \frac{B}{10^{12} \text{ G}} (1 + z)^{-1} \quad (1.24)$$

Therefore, the cyclotron line energy in an observed spectrum provides a direct estimation of the magnetic field of the neutron star.

1.9 Motivation of the thesis

High mass X-ray binaries are among the brightest X-ray sources harbouring a vast majority of accretion powered pulsars, i.e., highly magnetized accreting neutron stars. They display periodic and aperiodic variabilities on very diverse timescales from seconds to days or months. The most prominent periodic variability is the pulsation which is caused by the rotation of the neutron star. The orbital modulation is the variation in the X-ray intensity due to the orbital motion of the neutron star. This modulation can occur from the complete obscuration of the X-ray source by the companion (known as eclipse) or other orbital phase dependent factors like the absorption and scattering of X-rays in the stellar wind of the companion, accretion streams, etc. BeXRBs show a broad range of luminosity from their quiescent to outburst phases. These variabilities allow us to know the different physical conditions within the accretion column, accretion geometry, the structure and properties of the wind, the X-ray emission mechanisms at different luminosities, effects of the accretion torque, and its manifestation in spin period changes, etc.

The aim of this thesis is to gain new and further insights on HMXB pulsars by investigating their X-ray variability and accretion mechanism. We carried out X-ray timing and spectroscopic studies of individual sources which contain both supergiant and Be X-ray binaries. We studied the orbital intensity modulation of supergiant HMXB IGR J16393-4643 to probe the unique partial nature of its eclipse (chapter 3). We looked at the broadband X-ray characteristics of BeXRB GRO J2058+42 during its Type II outburst and quiescent. We also studied its accretion torque behaviour over a range of luminosity (chapter 4). We examined the pulse profiles of four persistent pulsars (Cen X-3, Vela X-1, GX 301-2 and OAO 1657-415) to find out any possible change in the accretion geometry at different accretion torque (chapter 5). The observation and analysis of these massive binaries are very useful in order to advance our present understanding of the X-ray emission mechanisms and the physics of accretion.

Chapter 2

**A description of the instruments and
data analysis techniques**



2.1 Introduction

We have used data from various observatories to achieve our scientific goals. In chapter 3, we have reported results from the archival data from the Burst Alert Telescope (BAT) and the X-Ray Telescope (XRT) aboard the Neil Gehrels Swift Observatory. In chapter 4, results from observations with NuSTAR, Fermi-GBM, Swift-BAT and MAXI-GSC have been reported. In chapter 5, we carried out our analysis by using the data from Fermi-GBM, which was obtained on request from the Fermi team. In this chapter, we will provide a brief description of the instruments and discuss the data analysis methods.

2.2 X-ray Instruments

2.2.1 Neil Gehrels *Swift* Observatory

Swift was launched on 20th November 2004 from Cape Canaveral Air Force Station, Florida. It is orbiting in a low earth orbit at an altitude of about 600 km with an inclination of 20.6 degrees. Swift has three scientific instruments onboard: A) the Burst Alert Telescope (BAT; [Barthelmy et al. 2005](#)), having an energy range of 15-150 keV; B) the X-Ray Telescope (XRT; [Burrows et al. 2005](#)), having an energy range of 0.2-10 keV, and the Ultra Violet Optical Telescope (UVOT; [Roming et al. 2005](#)), which has a UV and optical coverage of 170-600 nm. The three instruments are combinedly capable of rapid position determination of Gamma-Ray Bursts (GRBs) with an accuracy of arcseconds.

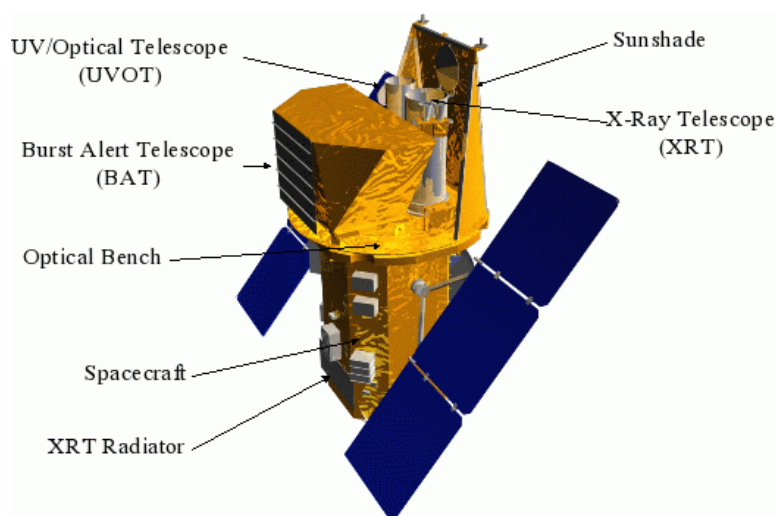


Figure 2.1: The Swift observatory with its three instruments. Image credit : <https://www.swift.ac.uk/about/instruments.php>.

2.2.1.1 Burst Alert Telescope

The Burst Alert Telescope (BAT) is a highly sensitive and large FoV telescope, primarily designed to detect Gamma-Ray Bursts. On detection of a burst, BAT calculates an initial position within ~ 10 seconds and sends the position to the spacecraft attitude control system. While observing bursts, BAT simultaneously surveys the hard X-ray sky. The detector plan consists of 2,768 pieces of $4 \times 4 \text{ mm}^2$ CdZnTe elements with 2 mm thickness. A coded-aperture mask is located at 1 meter above the detector plane. It is composed of 54000 pieces of $5 \times 5 \text{ mm}^2$ and 1 mm thick led tiles. The mask yields an area of 2.7 m^2 with a point spread function (PSF) of 17 arcmin and half-coded FOV of $100 \text{ degrees} \times 60 \text{ degrees}$, or 1.4 steradians. BAT operates in the burst mode and the survey mode. In the survey mode, it searches for new GRB trigger and collects hard X-ray survey data. In the burst mode, it locates the burst position. The BAT hard X-ray transient monitor provides continuous coverage ($80\% - 90\% \text{ day}^{-1}$) of the sky in the energy range of 15-50 keV (Krimm et al. 2013). The lightcurves of about a thousand of sources starting from 2005 February are publicly available and can be downloaded from the *Swift* website¹ in FITS or ASCII format. These lightcurves are generated on two time scales : single *Swift* pointing (continuous observation pointed at the same location) from 64s to 1200s and one-day averages. In chapter 3 and 4, we have used the daily averaged monitoring light curves of IGR J16393-4643 and GRO J2058+42 in 15-50 keV energy range.

2.2.1.2 X-Ray Telescope

The XRT is a focusing telescope with grazing incidence Wolter-I mirrors which focuses the incoming X-Ray onto an E2V CCD-22 detector. The telescope has a focal length of 3.5m, with a field of view of $23.6 \times 23.6 \text{ arcmin}$ and point spread function (PSF) of 18 arcsec at 1.5 keV. The overall instrument has a length of 4.67m and a diameter of 0.51m. The XRT mirror assembly is constructed with the X-ray mirror module (12 nested Wolter-I mirrors), thermal baffle, a mirror collar, and an electron deflector. The thermo-electrically cooled CCD-22 detector has an imaging array of 600×600 pixels (pixel size of $40 \times 40 \mu\text{m}$) and a storage region of 600×600 pixels (pixel size of $39 \times 12 \mu\text{m}$) and 2.5 arcsec/pixel resolution. It has four calibration sources located at each corner of the detector. The effective area is 110 cm^2 at 1.5 keV that decreases to 70 cm^2 at 8.1 keV.

The XRT uses four readout modes to observe the dynamic variability from GRB afterglows: 1) Imaging Mode (IM), 2) Photo-diode Mode (PD), 3) Windowed Timing Mode

¹<http://swift.gsfc.nasa.gov/docs/swift/results/transients/>

Table 2.1: Swift-BAT and XRT instrument characteristics

parameter	BAT	XRT
Energy	15–150 keV (imaging)	0.2 - 10 keV
Detector	CdZnTe (CZT)	e2v CCD-22
Effective area	5240 cm^2	120 cm^2 at 1.5keV
Detector operation	photon counting	photon-counting, integrated imaging, and timing
Field of view	1.4 sr (half-coded)	23.6 $\times$ 23.6 arcminutes
Detector elements	256 modules of 128 elements/module	600 $\times$ 602 pixels
Telescope PSF	22 arcminutes FWHM	18 arcseconds at 1.5 keV 22 arcseconds at 8.1 keV

(WT), and 4) Photon Counting (PC) Mode. The Imaging Mode produces a prompt integrated image to provide a rapid evaluation of the coordinates of the observed source. It does not produce suitable data for spectroscopy or time-resolved data. The Photo-diode Mode is designed to obtain data with high timing accuracy from rapidly varying very bright sources. The obtained data does not carry any spatial information but produces spectrum and lightcurve with a time resolution of 1.4 *ms*. This mode has been no longer available since May 2005. The Windowed Timing Mode carries one-dimensional position information with a high time resolution of 1.8 *ms*. The Photon Counting mode provides full spectral and spatial information with a time resolution of 2.5 *seconds*.

2.2.2 The Monitor of All Sky X-ray Image (*MAXI*)

The Monitor of All-Sky X-ray Image (*MAXI*) is a Japanese All-sky X-ray Monitor (ASM), scanning the whole sky since 2009. It was installed on the Japanese Experiment Module – Exposed Facility (JEM-EF or Kibo-EF) on the ISS (International Space Station) and was launched by space shuttle Endeavor ([Matsuoka et al. 2009](#)). *MAXI* was designed to achieve the best sensitivity and the highest energy resolution than any previous X-ray

all sky monitor payloads. It has two kinds of X-ray slit cameras, Gas Slit Camera (GSC, [Mihara et al. 2011](#)) and Solid-state Slit Camera (SSC, [Tomida et al. 2011](#)). These two cameras work in two complementary energy bands: GSC in the 2-20 keV band and SSC in the 0.7-10 keV band. We have used only GSC for our thesis work.

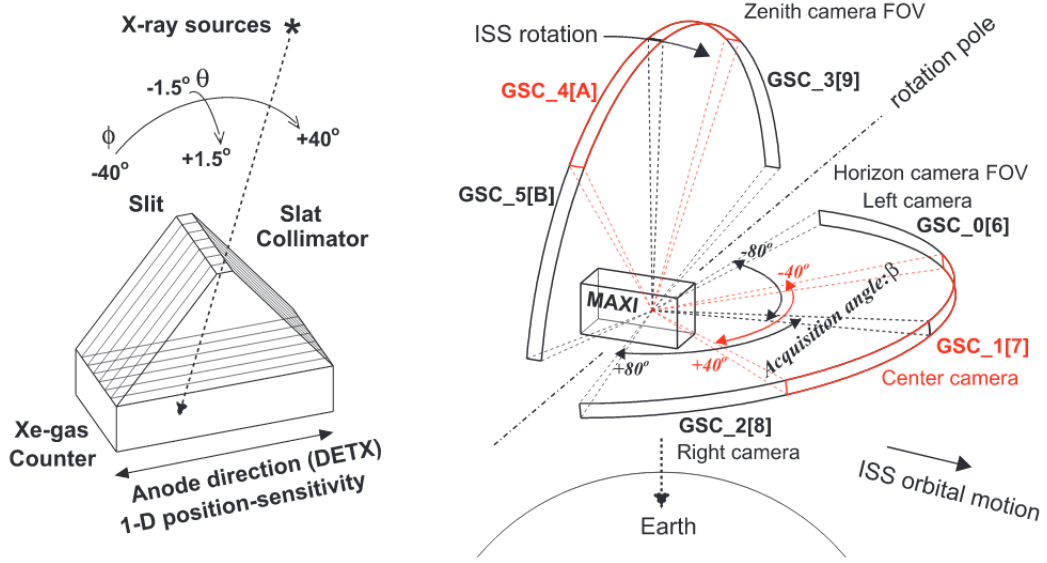


Figure 2.2: Left figure is a schematic diagram of a single GSC camera unit. Right figure is a schematic diagram of the GSC camera with the 12 counters (GSC_0, ..., GSC_9, GSC_A, GSC_B) on the ISS low-earth orbit ([Sugizaki et al. 2011](#)).

2.2.2.1 GSC

The Gas Slit Camera module is composed of a conventional slit camera with Xenon gas proportional counters and slit-and-slat collimators. The twelve proportional counters provide a total detector area of 5350 cm^2 and a FOV of 1.5° in FWHM. The slit-and-slat collimator has parallel tungsten rods with 3.7 mm separation, which constitutes the opening slit of the camera. The 0.1 mm thick phosphor bronze slats are aligned vertically to the slit rods. The entire GSC has six identical units to cover a collective field of view of $1.5 \times 160 \text{ deg}^2$. Each unit has one slit-and-slat collimator and two proportional counters, and the units are assembled into two modules, the horizon module GSC-H and the zenith module GSC-Z. They cover the earth's horizon and the zenith direction with equal area, which enables to cover about 85% of the sky at every ISS revolution (~ 92 minutes).

The proportional counters utilize resistive carbon-wire anodes that obtain one dimensional position sensitivity. The gas cell is separated into six carbon-anode cells for X-ray

detection and ten tungsten-anode cells for veto detector. The Xenon ($\sim 99\%$) gas is used with 1% CO_2 at a pressure of 1.4 *atm* at 0° C.

The visibility and the area of the GSC units continuously change according to the ISS motion. Thus, the time-dependent effective area should be known to derive the correct flux of the observed source. Effective areas are reconstructed by the lightcurve response builder. The lightcurve response builder assesses the ISS location, attitude, the source coordinates, and the GSC unit area as a function of the source incident angle. The energy response builder determines the effective area and the energy redistribution matrix for a given source and observation period. The specifications of the in-orbit performance of the MAXI-GSC are listed in Table 2.3.

Table 2.2: GSC in-orbit performance (Sugizaki et al. 2011)

parameter	specification
Spatial resolution	1.5 deg (FWHM)
Source localization accuracy	0.2 deg (90% containment radius)
All-sky coverage	85% per orbit (nominal 92 minutes), 95% per day
Exposure	4000 $cm^2 s$ per day (typical)
Energy resolution	18% (1σ) at 5.9 keV
Effective area accuracy	3% (1σ) at 4-10 keV
Sensitivity	15 mCrab per day in 2-30 keV (typical)

The *MAXI* data products (lightcurves and spectra) are publicly available in its website¹. GSC lightcurves with timescales of 1 day bin and 1 orbit (~ 90 min) bin can be downloaded in three (2-4 keV, 4-10 keV and 10-20 keV) and the total (2-20 keV) energy band since the beginning of the *MAXI* observation. We have used on-demand data from GSC to study the long term lightcurve of the Be X-ray binary GRO J2058+42 during the outburst in chapter 4.

¹<http://maxi.riken.jp/top/index.html>

2.2.3 The Nuclear Spectroscopic Telescope Array (NuSTAR)

The Nuclear Spectroscopic Telescope Array (*NuSTAR*) Small Explorer mission was launched on a Pegasus XL rocket on 13th June 2012. *NuSTAR* revolves in a low-Earth equatorial orbit (with inclination ~ 6 degrees) at an altitude of 600 km. It is the first focusing hard X-ray mission in orbit that detects X-rays in 3 to 78.4 keV energy band ([Harrison and NuSTAR Team 2013](#)). The *NuSTAR* optics and the detector are separated by a lightweight, extendable mast to achieve a focal length of 10 m. The X-ray optics have two co-aligned grazing incidence Wolter-I conical approximation mirror modules. Each mirror module consists of 133 nested multilayer-coated mirror shells. The two telescopes have two independent focal plane modules (FPMA and FPMB), each with four hybrid CdZnTe detectors. The detectors are arranged in 2×2 array shielded by CsI crystals. Each detector sizes about $2 \times 2 \times 0.2$ cm segmented into a 32×32 pixel array. The *NuSTAR* telescopes have $18''$ FWHM imaging resolution with a characteristic spectral resolution of 400 eV FWHM at 10 keV and a temporal resolution of $2 \mu\text{s}$.

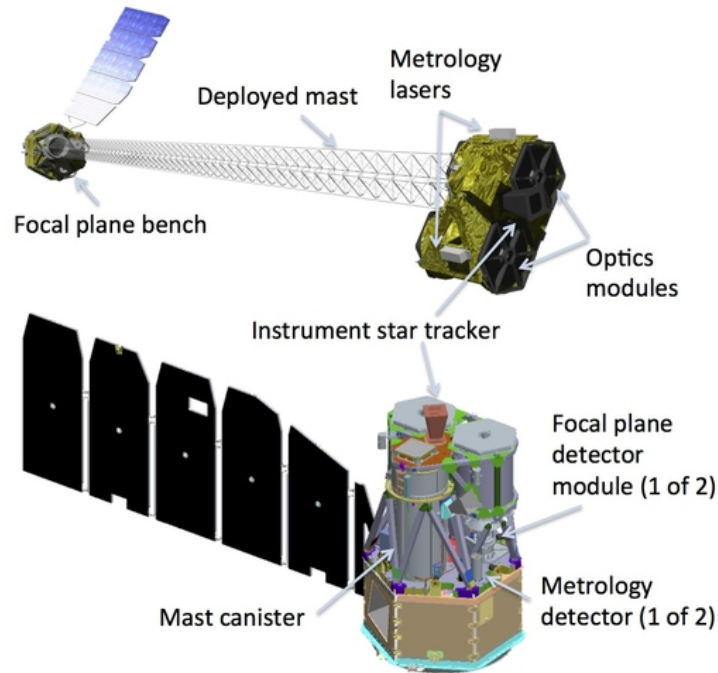


Figure 2.3: A schematic diagram of the observatory in the stowed (bottom) and deployed (top) configurations ([Harrison and NuSTAR Team 2013](#)).

Table 2.3: NuSTAR performance parameters.

parameter	specification
Energy	3 - 78 keV
Field of View	12.2×12.2 arcmin
Effective area accuracy	800 cm^{-2} at 6 keV 125 cm^{-2} at 60 keV
Angular resolution	58" (HPD), 18" (FWHM)
Energy resolution (FWHM)	400 eV at 10 keV 900 eV at 68 keV
Time resolution	2 μs
Detector sensitivity (in $erg\ cm^{-2}\ s^{-1}$)	2×10^{-15} (6-10 keV) 1×10^{-14} (10-30 keV)

2.2.4 The Fermi Gamma-Ray Space Telescope

The Fermi Gamma-Ray Space Telescope (FGST), initially known as the Gamma-ray Large Area Space Telescope (GLAST), was launched on June 11, 2008, from Cape Canaveral Air Force Station on board a Delta II Heavy rocket. It is placed in a low-Earth circular 96 minutes orbit with inclination of $\sim 26^\circ$ at an altitude of 560 kilometers. Fermi comprises two science instruments. The main instrument is the Large Area Telescope (LAT), a pair conversion telescope. LAT covers the energy range from 20 MeV to 300 GeV and observes a wide variety of Gamma-ray sources. The other instrument is the Fermi Gamma-Ray Burst Monitor (GBM), operating between 8 keV - 40 MeV. It complements the LAT by providing near-real time burst locations and performs the time-resolved spectroscopy of GRBs detected by both the GBM and the LAT. GBM also monitors X-ray pulsars using the earth occultation technique. We have used the GBM monitoring data of several pulsars in our study.

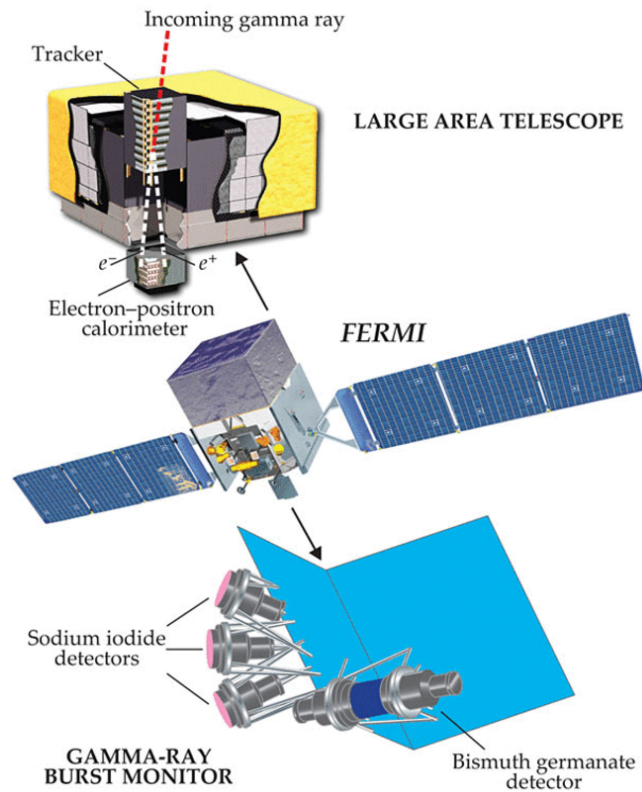


Figure 2.4: The Fermi Gamma-ray Space Telescope and its instruments ([Michelson et al. 2010](#)).

2.2.4.1 GBM

It is a hard X-ray all sky monitor consisting of multiple wide field-of-view detectors placed in different sides of the spacecraft in different orientations. These detectors are individually mounted around the spacecraft resulting in exposure of the whole sky without any occultation by the Earth. There are 12 thin Thallium activated Sodium Iodide crystal (NaI) detectors and 2 cylindrical Bismuth Germanate crystal ($\text{Bi}_4\text{Ge}_3\text{O}_{13}$ or BGO) detectors. The NaI detectors are composed of cylindrical crystal disks and sensitive in the energy range of 8 keV - 1 MeV. The crystals disks have a diameter and thickness of 1.27 cm and are packed in a hermetically sealed light-tight Aluminum housing. The two BGO detectors are placed on the opposite sides of the spacecraft and operate between 0.2 MeV to 40 MeV. These detectors are cylindrical crystals with a diameter and length of 12.7 cm. All the NaI and BGO detectors are coupled to the 10-stage 5-inch photo-multiplier tubes (PMT, which convert the scintillation light to an electric signal.

The data processing unit (DPU) of the GBM transmits three types of science data:

CTIME (continuous time data), CSPEC (continuous spectroscopy), and TTE (time trigger event) (Bissaldi et al. 2009). The CTIME data has a high time resolution of 256 ms (configurable from 64 ms to 1.024 s in units of 64 ms) and an energy resolution of 8 channels. The CSPEC data has a time resolution of 4.096 s (configurable from 1.024 s to 32.768 s in units of 1.024 s) and a high energy resolution of 128 channels. The TTE data comprises the individual detector events with the time resolution of 2 μ s and the same channel boundaries as CSPEC. In this thesis, we have used the CTIME data from GBM in chapters 4 and 5.

2.3 Data analysis techniques

Detection of an X-ray source happens when the X-ray photons reach the detector, and the information is stored in event files. We can have information about the position of the source in the sky, arrival time and the energy of the X-ray photons, and the brightness of the source. We have extracted and processed this information by using HEASOFT (High Energy Astrophysics Software). It includes packages like FTOOLS (general and mission-specific tools to manipulate FITS files) and XANADU, which consists of high-level, multi-mission tasks for spectral (XSPEC), timing (XRONOS), and imaging (XIMAGE) analysis. We have used them to do timing and spectral analysis in this thesis.

2.3.1 Timing analysis

Photometry is a fundamental technique in astronomy which refers to the measurement of the intensity or brightness of a celestial object at a given wavelength. After successful applications in the Optical band, it is now exercised through out the electromagnetic spectrum including Radio, Infrared, X-ray and so on. The X-ray band suffers from low photon counts as compared to the Optical band. Even the bright sources typically have few hundreds of photon counts. Therefore broadband photometry becomes more important in X-ray astronomy and is used to estimate flux, hardness ratio, and colors. While measuring the flux, the time resolution is crucial. We witness a wide range of periodic and aperiodic timing characteristics from the X-ray sources. Measuring the flux with a time resolution of minutes becomes irrelevant when observing a variability over seconds. On the other hand, a time resolution of seconds is ineffective if we want to detect the variation of source intensity over days. Timing analysis implies processing of the timing information of the detected photons. For this purpose, we have extracted lightcurves

using XSELECT from the event files. A lightcurve is defined by the photon count rates over a range of time. We have used the following tools to carry out our studies.

2.3.1.1 lcurve

We have used this tool to create binned lightcurves from the event data files. It uses the FITS file format for both input and output and plots up to 4 lightcurves at a time. *lcurve* can produce the ratio of the count rates from two lightcurves at a given time duration. It is called the ‘Hardness ratio’.

2.3.1.2 lcmath

It can produce the difference or the sum of the count rates of two lightcurves. The difference is used for subtracting the background photon counts, and the sum can be used to create a combined lightcurve from two detectors.

2.3.1.3 efsearch

This tool is used to search for periodicities in the lightcurves. It folds the time series over a range of periods and searches for the maximum χ^2 as a function of period. It produces a χ^2 versus period plot, and the maximum χ^2 corresponds to the most likely period. We determined the orbital period in chapter 3 and pulse periods in chapter 4 using this tool.

2.3.1.4 efold

This tool is used for folding or stacking the lightcurve with a period w.r.t. an epoch. It can plot up to four lightcurves simultaneously, and the standard output is normalized counts/sec versus phase. It also calculates the ratio and sum of the lightcurves. After averaging the folded lightcurves in a frame, the error bars are calculated either using the standard deviation of the mean values of each phase bin or propagating the error in each phase. In this thesis, *efold* has been extensively used to produce pulse profiles and orbital intensity profiles.

2.3.2 Spectral analysis

In astronomy, spectroscopy is used as a fundamental tool to derive physical parameters like temperature, chemical composition, magnetic field etc. which provide an insight to the physical processes in the celestial objects. It is the technique of splitting the electromagnetic radiation into its constituent wavelengths or energy. Depending on the wavelength or the energy of the photon, the interaction between the radiation and the matter can be very different for each waveband and therefore the spectroscopic techniques also

vary. X-ray photons have much higher energy and much shorter wavelengths than the Radio, Infrared or Optical band. These high energy photons can interact with individual electrons, atoms or molecules and can produce large amount of ionization. Therefore unlike the Radio, Infrared or Optical band, it is possible to measure individual photon energies in the X-ray band.

The X-ray detectors are designed primarily to detect energy (occasionally wavelength) of ionizing radiation. In this thesis, we have analysed observations from different types of detectors which use a variety of techniques like proportional counters (*MAXI-GSC*), CCDs (*Swift-XRT*), CdZnTe detectors (*NuSTAR* and *Swift-BAT*), scintillation detectors (*Fermi-GBM*). A proportional counter detects X-ray photon energy using an ionization chamber with a gas filled container. It operates in the same way as the Geiger-Muller counter but with a lower voltage that maintains a linear relation between the incoming photon energy and the size of the electric pulse output. To measure very high photon energies, a proportional counter must be build at high pressure or the size of the gas container should be enough to provide sufficient column density to the incident photon. The charge-coupled devices or the CCD detectors are being widely used on currently operating X-ray missions (like *Chandra-ACIS*, *XMM-EPIC* and *RGS*). They are array of linked capacitors that produce electron-hole pairs by absorbing X-ray photons. CCDs provide best X-ray sensitivity in the energy range of few hundred eV to 10 keV. Photons with lower energies are likely to get absorbed by any structure on the CCD surface and the higher energy photons are likely to pass through the active region of the detector without any interaction. CZT (or CdZnTe) detector can be used to detect photons with energy higher than 10 keV which has higher cross section than silicon and therefore a better quantum efficiency. CZTs can provide energy resolution of few percent within the energy range of 10-100 keV. Scintillator detectors are also used to detect high energy X-ray photons. They consist crystals (for example NaI and BGO) that scintillates and re-emits the absorbed X-ray photon energy in the form of optical light. There are several other type of detectors like microchannel plates, microcalorimeters (high energy resolution) etc. Although these X-ray detectors are primarily characterised by their energy range and energy resolution, the response of any X-ray detector to monochromatic X-rays is often non-linear, asymmetric, and there can be fluorescence photon loss from a detector causing escape peak etc.

We took spectroscopic data from *Swift-XRT* and *NuSTAR* and used XSPEC or the “X-ray Spectral Fitting Package” for the spectral analysis. A spectrum is detected as the distribution of the photon counts over the energy channels (I). The intrinsic spectrum

$f(E)$ of an X-ray source is related to the counts per channel $C(I)$ by

$$C(I) = \int_0^{\infty} f(E)R(I,E)A(E)dE \quad (2.1)$$

where $R(I,E)$ is the instrument response for the particular detector. It is a measure of the probability of an incoming photon of energy E to be detected in channel I . $A(E)$ is the effective area of the detector.

The spectral analysis is done by choosing a model spectrum $M(E)$ and fitting it to the observed spectra. Once the observed data is fitted with a predicted model, the best fit parameters and the errors are estimated. Then we check if the model and the parameters provide a good fit to the observed data. The goodness of fit is evaluated by using statistical tests that determine whether the data and the model with the best fit parameters are consistent with each other. The χ^2 fit statistics has been used in this thesis. It is given by,

$$\chi^2 = \frac{\sum (C(I) - M(I))^2}{(\sigma(I))^2} \quad (2.2)$$

$\sigma(I)$ is the error associated with each count $C(I)$. The model with minimum χ^2 is the best fit spectral model. Ideally, the reduced χ^2 is approximately 1 for the best fit model where the reduced χ^2 is defined by the χ^2 divided by the degrees of freedom. The photons arriving from an astronomical source result in a Poisson distribution. With large number of counts, the Poisson distribution can be approximated to the Gaussian (or normal) distribution. The χ^2 statistics is applicable for a data set with Gaussian distribution. For faint sources with low number of counts, the C-statistic (or *cstat*) is more suitable ([Cash 1979](#)).

2.3.2.1 Continuum

The continuum models of X-ray binary pulsars are typically described by a *power law* with an exponential cutoff in the range of 5-30 keV ([Bildsten et al. 1997](#) and references therein). The simple *power law* spectrum results from the Comptonization of emitted X-ray photons and is given by

$$powerlaw(E) = AE^{-\Gamma} \quad (2.3)$$

where E is the photon energy in *keV* and Γ is the photon index. The normalization

A is the photon flux at 1 keV in the unit of *Photons keV⁻¹ cm⁻² s⁻¹*. The most commonly used continuum model for X-ray pulsars is power law spectra with high-energy cutoff or HIGHECUT (White et al. 1983). Its analytical form is

$$HIGHECUT(E) = AE^{-\Gamma} \times \begin{cases} 1 & (E \leq E_{cut}) \\ e^{-(E-E_{cut})/E_{fold}} & (E > E_{cut}) \end{cases} \quad (2.4)$$

where E_{cut} and E_{fold} are the cutoff energy and folding energy. The cutoff energy indicates where the spectrum turns from regular power law to a steeper model. Tanaka (1986) suggested the Fermi-Dirac cutoff (FDCUT) model, which is very similar to the HIGHECUT model. It describes the spectrum based on the Fermi-Dirac distribution and is given by

$$FDCUT(E) = AE^{-\Gamma} \times \frac{1}{1 + e^{(E-E_{cut})/E_{fold}}} \quad (2.5)$$

Another frequently used continuum model is NewHCUT (Burderi et al. 2000). It is also a power law with an exponential cutoff smoothened around the cutoff energy. The analytical form is

$$NewHCUT(E) = \begin{cases} NE^{-\Gamma} & (E \leq E_{cut} - \Delta E) \\ AE^3 + BE^2 + CE + D & (E_{cut} - \Delta E < E < E_{cut} + \Delta E) \\ NE^{-\Gamma} e^{-(E-E_{cut})/E_{fold}} & (E \geq E_{cut} + \Delta E) \end{cases} \quad (2.6)$$

where A , B , C , and D are the constants computed under the assumption that the functional form of NewHCUT(E) is continuously derivable.

2.3.2.2 Photoelectric absorption

The intrinsic radiation from the X-ray source gets absorbed by the interstellar medium before reaching the detector. In the X-ray spectra, the absorption effect is visible in the lower energies (below 5 keV), whereas the hard X-ray spectra do not change much with increased absorption. Therefore, the soft X-ray spectra of high mass X-ray binaries have low photon counts due to high absorption. There are several models available in XSPEC for the absorption column density. *phabs* is the photoelectric absorption model. This multiplicative model is widely used to describe the continuum absorption by the interstellar medium. The analytical form of *phabs* is

$$M(E) = e^{-N_H \sigma(E)} \quad (2.7)$$

where $\sigma(E)$ is the photo-electric cross-section for the absorption of photon with energy E and N_H is the equivalent hydrogen column density (in units of $10^{22} \text{ atoms cm}^{-2}$).

2.3.2.3 Gaussian absorption line

The cyclotron resonance scattering feature or the cyclotron line is an absorption feature in the spectra of HMXB pulsars. It can be described by a Gaussian absorption profile or *gabs*, which is a multiplicative component in XSPEC. The analytical form of this component is

$$gabs(E) = e^{(-d/\sqrt{2\pi}\sigma)e^{-(E-E_{cyc})^2/2\sigma^2}} \quad (2.8)$$

where E_{cyc} , σ and d are the cyclotron line energy, width and depth. The optical depth of the line center is $\frac{d}{\sqrt{2\pi}\sigma}$.

2.3.2.4 Gaussian emission line

The strong iron fluorescence lines are often present in the spectra of XRBs. These emission lines are modeled with an additive Gaussian line component in XSPEC. It is

$$gauss(E) = \frac{A}{\sigma\sqrt{2\pi}} e^{-(E-E_l)^2/2\sigma^2} \quad (2.9)$$

where E_l and σ are the line energy and width in *keV*. A is the total *photons cm⁻² s⁻¹* in the line.

Chapter 3

The Partially Eclipsing High Mass X-ray Binary IGR J16393-4643



Based on:

Investigating a unique partial eclipse in the High Mass X-ray Binary IGR J16393-4643 with Swift-XRT

Kabiraj S., Islam N., Paul B., MNRAS **491**, 1491, 2020

3.1 Introduction

In an X-ray binary system, the X-ray emitting compact star can get occulted by its companion in every orbit and the X-ray intensity observed from the binary is reduced by a large factor. Variation in the orbital intensity depends upon the inclination of the orbital plane of the X-ray binary with respect to our line of sight, and the angular size of the companion star to the compact star. For compact emission regions, the eclipse is expected to be complete, except any scattered radiation from the environment.

For Low Mass X-ray Binaries (LMXBs), eclipses are predicted for a narrow range of inclination as the companion star usually has a smaller angular diameter to the compact star (Frank et al. 1987). The only exception to this can be the Accretion Disk Corona (ADC) sources among the LMXBs like 4U 1822-37, which shows a partial eclipse as the ADC is comparable in size to the companion star (Jain et al. 2010). Accretion disk corona was first used to explain narrow partial eclipses (White and Holt 1982). The X-rays heats up the gas in the corona of the compact star, which gives rise to residual flux which can be observed even during the eclipses. The presence of a corona can cause scattering of photons which leads to broad and shallow dips in the X-ray light curve. We also see low X-ray intensity states prior to the eclipse in some High Mass X-ray binaries (HMXBs) due to the presence of enhanced stellar wind in certain orbital phases, which lead to an orbital phase dependent absorption.

In an HMXB system with a neutron star, the X-ray emission region sizes about a few kilometers, whereas the typical velocity of a neutron star in such system is of the order of 200 km/s and the companion star is many orders of magnitude larger in size compared to the X-ray emission region. Therefore, X-ray eclipses are expected to be sharp and complete (Falanga et al. 2015). The eclipse ingress and egress are however gradual in soft X-rays due to strong absorption in the stellar wind which has large column density during the eclipse ingress and egress. However, hard X-ray eclipses are always known to be sharp. IGR J16393-4643 is an unusual HMXB system with a neutron star that shows a gradual and partial eclipse even in the hard X-rays (Islam et al. 2015).

3.2 IGR J16393-4643

IGR J16393-4643 was first discovered by *Advanced Satellite for Cosmology and Astrophysics* (ASCA) observatory and was named as AX J163904-4642 (Sugizaki et al. 2001). It

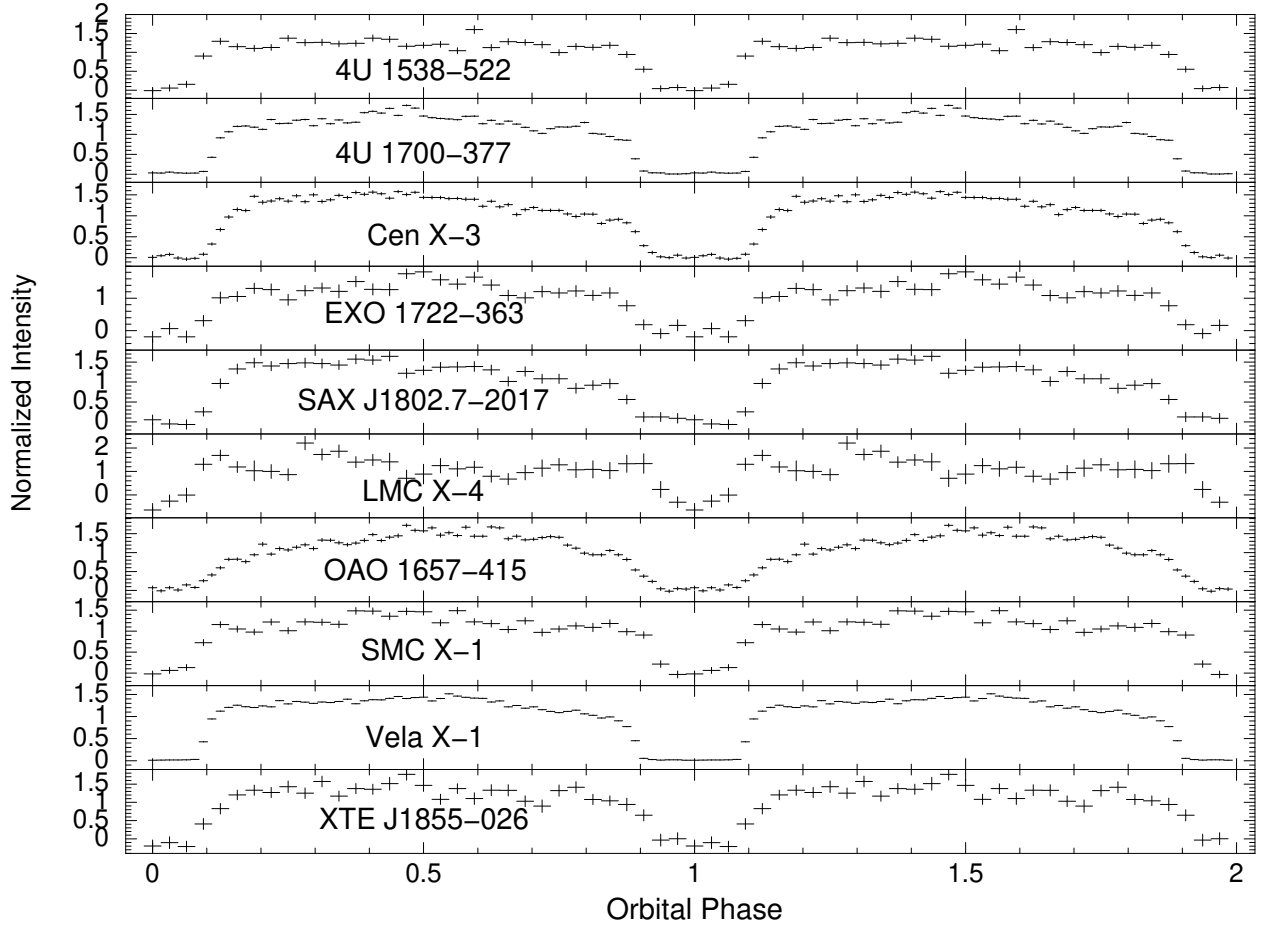


Figure 3.1: The hard X-ray orbital intensity profile of 10 HMXBs obtained using *Swift*-BAT long term light curves. The intensity in the eclipse goes to nearly zero in all these HMXBs.

was detected again in the *INTEGRAL* survey of the galactic plane (Bird et al. 2004). This X-ray binary system hosts a neutron star with a spin period of ~ 910 seconds (Bodaghee et al. 2006). The persistent pulsation and spectral characteristics suggests that it is a heavily absorbed wind accreting HMXB with an OB type supergiant counterpart. The presence of iron emission lines at 6.4 keV and 7.1 keV in its spectra was confirmed from *XMM-Newton* and *NuSTAR* observations. A cyclotron line at 29.3 keV indicates that the pulsar has a strong magnetic field of about 10^{12} Gauss (Bodaghee et al. 2016). However, the exact nature of the companion of the pulsar remained ambiguous. An infrared candidate, 2MASS J16390535-4642137 was suggested to be the companion of the pulsar. Later Bodaghee et al. (2012) found that the real companion is a distant reddened star, blended with the bright 2MASS star. With this assumption, it is either a supergiant O9 star with $R \sim 20R_{\odot}$ or a main sequence B star with $R \sim 10R_{\odot}$. If it is an O9 star with then the estimated distance is 25 kpc, which is probably too far to be the possible candidate. If it is a main sequence B star, the estimated distance is 12 kpc, which is physically reasonable due to the presence of a massive star-forming H II region at a similar distance (Russeil 2003).

Corbet et al. (2010) reported a significant modulation of its X-ray intensity with the Swift-BAT and RXTE lightcurve in every 4.24 days and it was inferred to be an “Eclipse” with an eclipse semi-angle $\sim 17^{\circ}$ (Islam et al. 2015). This indicated that the companion is either at an orbital inclination of 39° - 57° for radius with $R \sim 20R_{\odot}$ or at 60° - 77° for a radius with $R \sim 10R_{\odot}$. Interestingly, the eclipse profile of IGR J16393-4643 shows a partial eclipse of the neutron star in the *Swift*-BAT hard X-ray energy band, unlike any other eclipsing HMXBs. There is approximately 25% amplitude during the eclipse phase in comparison to the mean flux. The partial eclipse of IGR J16393-4643 makes it a unique candidate in eclipsing HMXB family which makes us intrigued about its true nature. It can be an actual eclipse by its companion or a dip indicating a periodic increase in the absorption column density.

3.3 Observation and Analysis

To unravel the mystery of this partial eclipse we aimed to determine the absorption column density (N_H) and the equivalent width of the Fe-K α line inside the partial eclipse state. A high value of N_H during the partial eclipse will indicate that the periodical drop in the intensity profile is not due to the occultation by its companion. But a moderate

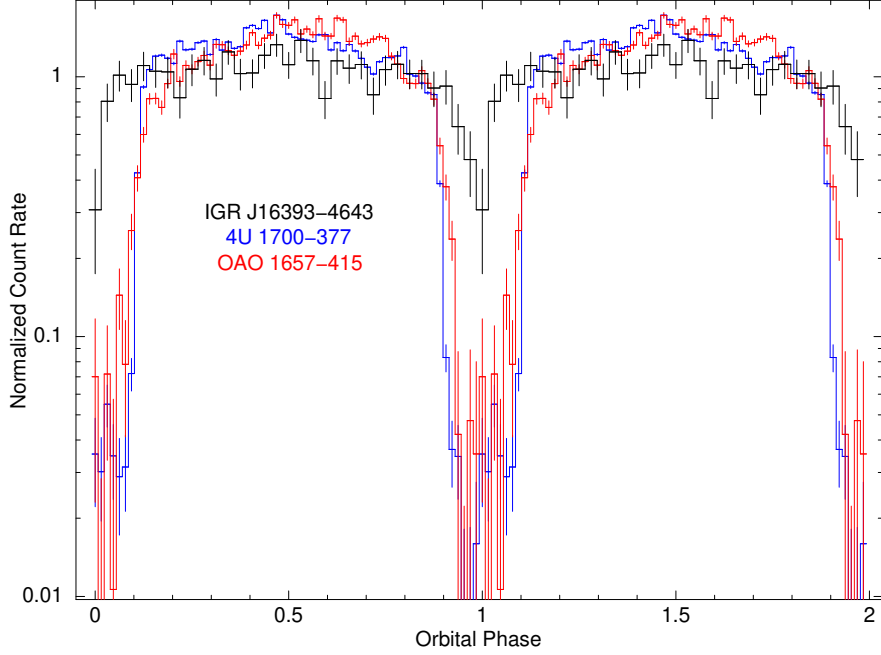


Figure 3.2: This plot is the overlaid orbital profiles of two of the HMXBs (4U 1700-37 and OAO 1657-415) from 3.1 and IGR J16393-4643 to show the difference of this source with the other eclipsing HMXBs. The orbital profile of IGR J16393-4643 is narrower and shallower indicating the partial eclipse of the X-ray source.

value of N_H in comparison to the out of eclipse state along with a substantial value of the equivalent width of the Fe-K α line (Aftab et al. 2019) will establish that it is actually an eclipse.

There were several observations by *Swift*-XRT taken in Photon Counting mode at different orbital phases from MJD 57411 to MJD 57554 within 0.2-10 keV energy band. The Photon Counting mode has a time resolution of 2.5 seconds and it provides full imaging and spectroscopic resolution. We selected 73 observations which were taken within 2 arcmin radius centering the source having exposure time more than 200 seconds. We used HEASOFT 6.22. to analyse these observations and tabulated the observation ID, exposure time, photon counts and orbital phase of all of them in Table 3.1. Minimum exposure time among these observations was 211 seconds and the maximum time is 1785 seconds, and the lowest and highest count-rate was 0.02 count/s and 0.39 count/s. From the cleaned event files, we extracted the source photons from a region of 60'' radius centering the source. We extracted the background photons from a source free region with 60'' radius within the FoV, where no other X-ray source was present. We have generated

lightcurves and spectra from source and background region files using XSELECT v2.4d.

We generated the exposure maps and used those to create the ancillary response file (ARF). The required response matrix file (RMF) were obtained from the calibration database CALDB 1.0.2. We have generated the spectra for source and background for each observation and used the ARF and RMF files to get the required spectra for analysis. Therefore, in total we had 73 lightcurves and spectra to do the timing and spectral analysis.

3.3.1 Orbital intensity profile

We analyzed the 15-50 keV *Swift*-BAT long term lightcurve (MJD 53415.99 to MJD 58712.52) and did a period search using ‘efsearch’. The time between the start and end time for the used BAT lightcurve is 5296 days and the effective total exposure is 45 Ms. We found a period of 366106 seconds (~ 4.24 days) with the epoch of MJD 53418.3 so that the minimum intensity is positioned at phase zero. We folded the lightcurve with the same period and epoch. We have used 32 time bins per period, same as in [Corbet and Krimm \(2013\)](#) and [Islam et al. \(2015\)](#) due to limited statistics in this source. The eclipse duration is for 3-4 bins.

We have extracted the lightcurves with a default bin time of 2.5 seconds from each of the XRT observations and folded them with time period of 366106 seconds (i.e. ~ 4.24 days) and with the epoch of MJD 53418.3. These observations are almost evenly spread providing us full coverage of the entire orbital phase. The duration of the low intensity state is of about 1/5th of the total orbital period. The top and bottom panel of [Figure 3.3](#) shows the modulation in the orbital intensity profile from *Swift*-BAT and *Swift*-XRT observations respectively. The intensity fell to $\sim 30\%$ inside the eclipse phase indicating it to be partial in soft X-rays.

[Figure 3.3](#) represents both the lightcurves from BAT (top panel) and XRT (bottom panel). The eclipse profile is narrow in both soft and hard X-rays. The solid, dashed and dotted lines enclose the region of phase 0.91-1.09, 0.31-0.33 and 0.59-0.61 respectively which we have used for spectral analysis.

Table 3.1: Log of all observations with exposure time, total number of source photons, average count-rate and orbital phase.

Order No.	Observation MJD	Observation ID	Exposure time (seconds)	Total no. of photons (counts)	Average count-rate (counts/sec)	Orbital Phase
1.	57411	00034135004	1056	117	0.11	0.31
2.	57413	00034135006	939	89	0.09	0.81
3.	57417	00034135010	966	118	0.12	0.74
4.	57418	00034135011	929	57	0.06	0.93
5.	57420	00034135013	1011	134	0.13	0.59
6.	57422	00034135015	1785	115	0.06	0.98
7.	57423	00034135016	986	252	0.25	0.15
8.	57424	00034135017	864	44	0.05	0.56
9.	57425	00034135018	1064	114	0.11	0.73
10.	57426	00034135019	999	15	0.02	0.01
11.	57427	00034135020	852	91	0.11	0.23
12.	57428	00034135021	941	111	0.12	0.31
13.	57430	00034135023	1321	168	0.13	0.78
14.	57431	00034135024	996	34	0.03	0.07
15.	57432	00034135025	779	81	0.10	0.24
16.	57433	00034135026	847	75	0.09	0.52
17.	57434	00034135027	864	101	0.12	0.78
18.	57435	00034135028	859	93	0.11	0.12
19.	57437	00034135030	1069	158	0.15	0.59
20.	57438	00034135031	1046	82	0.08	0.82
21.	57439	00034135032	1064	42	0.04	0.06
22.	57441	00034135034	1016	121	0.12	0.51
23.	57442	00034135035	1056	121	0.11	0.60
24.	57443	00034135036	859	42	0.05	0.98
25.	57444	00034135037	924	78	0.08	0.15
26.	57445	00034135038	217	22	0.10	0.34
27.	57446	00034135039	1346	150	0.11	0.61
28.	57448	00034135041	946	98	0.10	0.14
29.	57458	00034135051	1054	95	0.09	0.45
30.	57459	00034135052	969	91	0.09	0.65
31.	57462	00034135055	984	169	0.17	0.50
32.	57464	00034135057	946	32	0.03	0.92
33.	57466	00034135059	1021	101	0.10	0.33

Order No.	Observation MJD	Observation ID	Exposure time (seconds)	Total no. of photons (counts)	Average count-rate (counts/sec)	Orbital Phase
34.	57469	00034135062	817	75	0.09	0.10
35.	57471	00034135064	330	55	0.17	0.63
36.	57473	00034135066	789	25	0.03	0.07
37.	57479	00034135072	1039	107	0.10	0.50
38.	57481	00034135074	916	70	0.08	0.95
39.	57483	00034135076	1031	119	0.12	0.44
40.	57484	00034135077	944	90	0.10	0.72
41.	57485	00034135078	897	352	0.39	0.74
42.	57486	00034135079	944	38	0.04	0.98
43.	57490	00034135083	844	59	0.07	0.07
44.	57491	00034135084	986	152	0.15	0.26
45.	57492	00034135085	954	141	0.15	0.60
46.	57493	00034135086	971	148	0.15	0.62
47.	57496	00034135089	1074	179	0.17	0.50
48.	57497	00034135090	976	103	0.11	0.78
49.	57498	00034135091	912	70	0.08	0.89
50.	57499	00034135092	594	24	0.04	0.06
51.	57503	00034135096	1091	84	0.08	0.07
52.	57504	00034135097	996	96	0.10	0.38
53.	57506	00034135099	1036	78	0.08	0.83
54.	57508	00034135101	1044	122	0.12	0.29
55.	57509	00034135102	1039	93	0.09	0.53
56.	57510	00034135103	922	162	0.18	0.79
57.	57514	00034135107	894	126	0.14	0.64
58.	57516	00034135109	917	39	0.04	0.19
59.	57519	00034135112	587	27	0.05	0.91
60.	57524	00034135117	1044	32	0.03	0.01
61.	57525	00034135118	922	100	0.11	0.32
62.	57527	00034135120	211	96	0.45	0.79
63.	57528	00034135121	1126	119	0.11	0.89
64.	57530	00034135123	862	47	0.05	0.47
65.	57531	00034135124	912	80	0.09	0.69
66.	57532	00034135125	979	18	0.02	0.02
67.	57533	00034135126	862	194	0.23	0.25
68.	57538	00034135131	989	107	0.11	0.32

Order No.	Observation MJD	Observation ID	Exposure time (seconds)	Total no. of photons (counts)	Average count-rate (counts/sec)	Orbital Phase
69.	57549	00034135142	954	26	0.03	0.01
70.	57551	00034135144	969	231	0.24	0.47
71.	57552	00034135145	1031	189	0.18	0.70
72.	57553	00034135146	1036	37	0.04	0.92
73.	57554	00034135147	1019	152	0.15	0.14

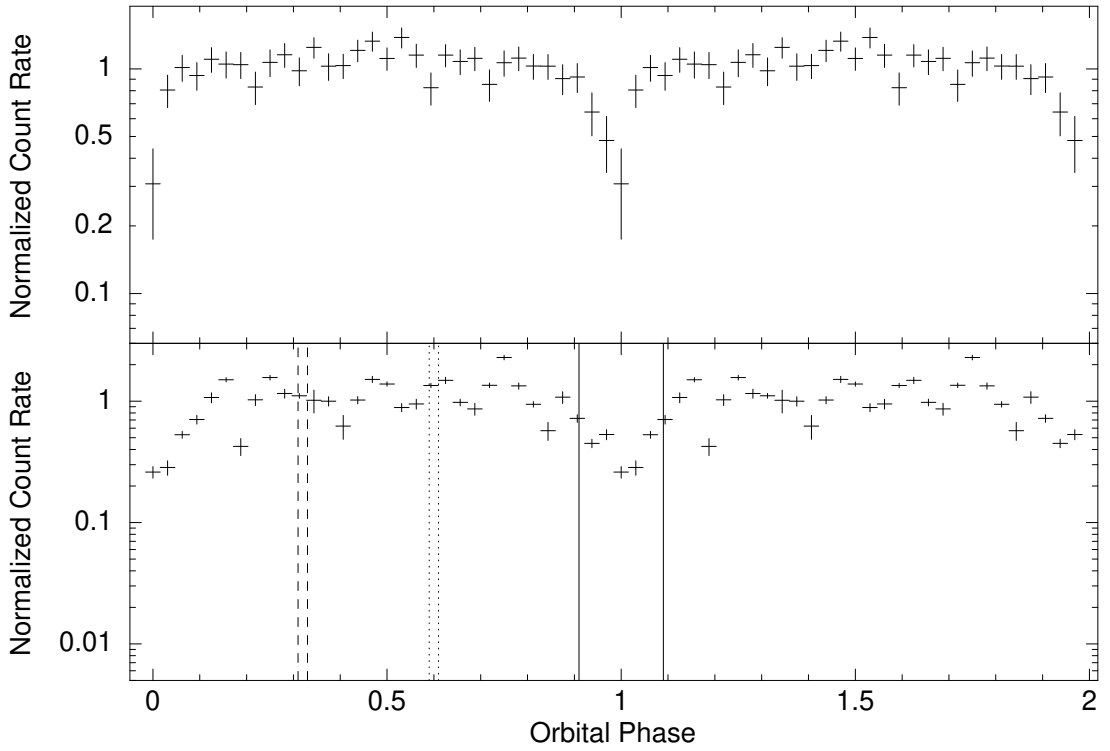


Figure 3.3: Top panel: The intensity profile of HMXB IGR J16393-4643 constructed with *Swift*-BAT lightcurve folded with a period of 366106 seconds with the epoch of MJD 53418.300 to have the minimum counts at phase 1. The low intensity state appears to be an eclipse.

Bottom panel: *Swift*-XRT light curves, folded with the same period. The dashed and dotted lines enclose the phase 0.31-0.33 and 0.59-0.61 respectively and the region between the two solid lines (i.e. phase 0.91-1.09) indicates the partial eclipse phase.

3.3.2 Spectral analysis

This partial eclipse profile, both in hard and soft X-rays, makes us inquire about the spectral nature of it. Individual XRT observations were too short for spectral analysis. We have therefore combined multiple observations in certain phase ranges for spectral analysis. There were 17 observations which are within the phase range of the partial eclipse (0.91 to 1.09). In Figure 3.3, these 17 observations fall in between the two solid lines. To make a comparison of the eclipse spectrum with spectrum during out of eclipse we have combined four out of eclipse observations in phase range 0.31-0.33 and four observations in phase range 0.59-0.61. Table 3.1 shows that both the phase range 0.31-0.33 and 0.59-0.61 have at least four observations, whereas there are fewer number of observations at other phases. Therefore, we choose these two particular phases to add up the spectra that can result into an improved statistic. The ranges of the out of eclipse spectra are also denoted in Figure 3.3 with pair of dashed and dotted lines respectively. We have merged the ARF and RMF file separately for each observation.

Then to get a particular added spectra, we have added the source and background spectra separately and added merged ARF-RMF files and combined them to get a background subtracted source spectra with RMF. The total exposure time for the added spectra during the partial eclipse phase is ~ 15 kilosec, and during phase 0.31-0.33 and 0.59-0.61, it is ~ 4 kilosec.

We have fitted all the three X-ray spectra at different orbital phases within the energy range 2.0-9.0 keV using XSPEC v12.9.1m. These spectra were modeled with a power law and photoelectric absorption due to the absorbing matter along our line of sight. The absorption column density, photon index and the total flux of these three spectra have been tabulated in Table 3.2. We do not notice any significant change in the column density inside and outside the eclipse. All of these spectra are in the energy range of 2.0-9.0 keV and are shown in Figure 3.4. We have overlaid the eclipse spectrum with the out of eclipse spectrum at the two phase ranges 0.31-0.33 and 0.59-0.61 (Figure 3.5) and we see the spectral profiles are very similar inside and outside the low intensity state. Overall, the eclipse spectrum is lower by a factor of about 2-3.

For spectra with limited statistics, there is often dependence between the column density and photon index. To examine the dependence and to check the goodness-of-fit

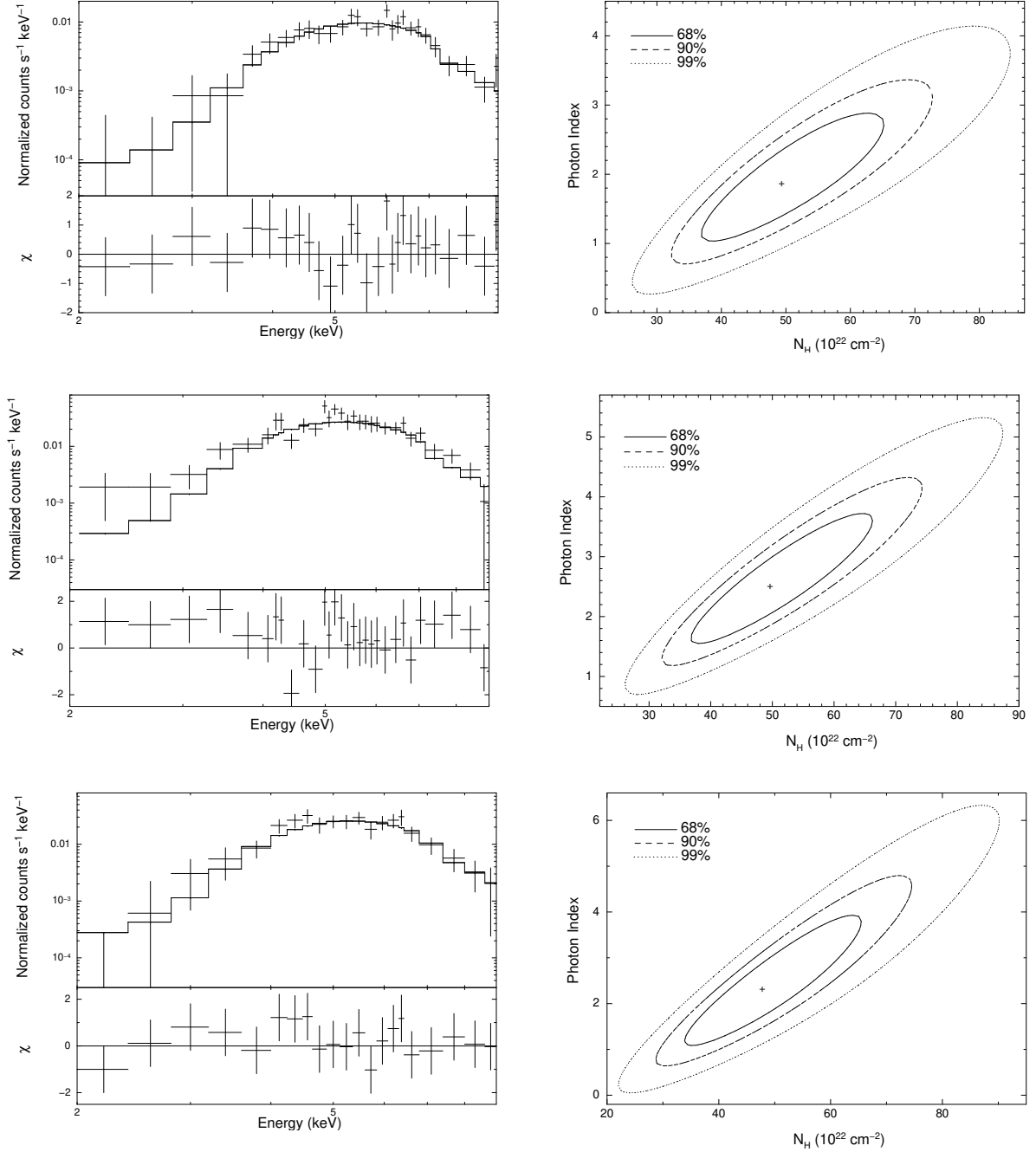


Figure 3.4: Left column: The added X-ray spectrum from the observations during the partial eclipse phase (at the top), during phase 0.31-0.33 (in the middle) and during phase 0.59-0.61 (at the bottom) fitted with an absorbed power-law model. Right column: The contour plot of χ^2 between column density and photon index for each spectra.

Table 3.2: Spectral parameters for added Swift-XRT observations with quoted for 90% confidence limits.

Phase	Column Density N_H^a	Photon Index (Γ)	Total Flux ^b	Upper limit of EQW^c Fe K α line (keV)	$\chi^2/d.o.f.^d$
0.91-1.06	49^{+17}_{-14}	$1.9^{+1.1}_{-0.9}$	$8.28^{+0.41}_{-6.93}$	0.346	0.729
0.27-0.29	50^{+18}_{-14}	$2.5^{+1.3}_{-1.1}$	$18.10^{+0.71}_{-10.03}$	0.262	1.004
0.55-0.57	48^{+16}_{-15}	$2.3^{+1.8}_{-1.3}$	$21.32^{+0.17}_{-14.12}$	0.345	0.529

^a : units of 10^{22} atoms cm^{-2}

^b : units of 10^{-12} ergs cm^{-2} sec^{-1} & in the 2-9 keV range

^c : equivalent width (EQW)

^d : without a gaussian component at 6.4 keV

to the model, we have used the chi-square statistic. It provided the best fit value for column density and photon index for each of the three spectra. In the left column of Figure 3.4, there are three χ^2 contour plots. The “+” sign is where the χ^2 is minimum, and 68%, 90% and 99% confidence contours are plotted with the solid, dashed and dotted lines.

An iron emission line at 6.4 KeV was detected in this source with the XMM EPIC-PN by Bodaghee et al. (2006). Also, 6.4 keV lines are ubiquitous in HMXBs (Giménez-García et al. 2015). However, in all the three spectra at different phases, we do not observe any significant presence of the iron emission line. For a possible presence of a line at 6.4 keV, we have estimated the upper limit of the equivalent widths at 90% confidence level in Table 3.2.

3.4 Results and Discussion

In this chapter, we have studied the periodic variability of the X-ray emission from the HMXB IGR J16393-4643 corresponding to its orbital phase to understand the nature of

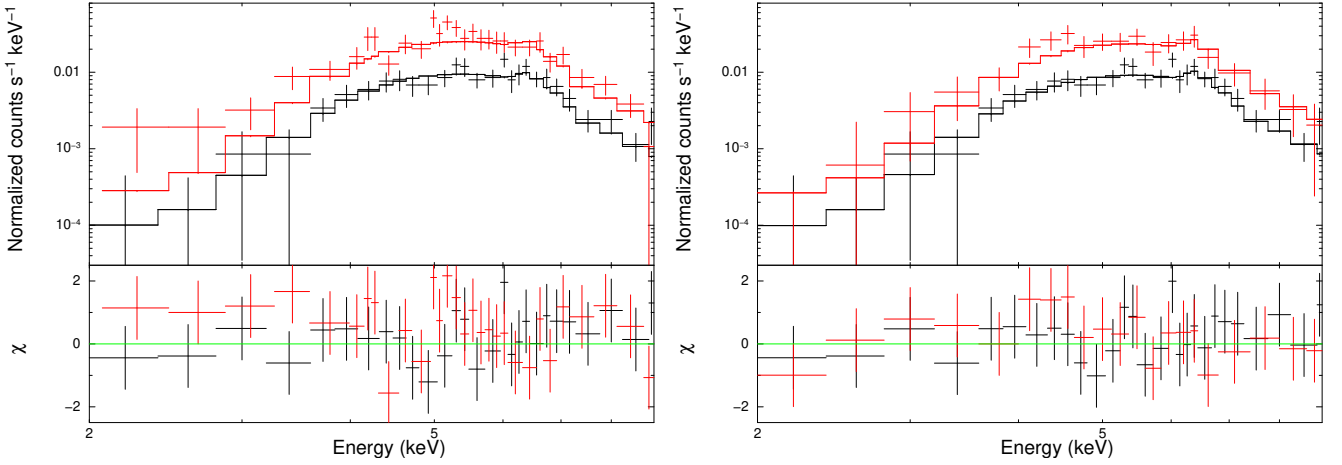


Figure 3.5: The overlaid added X-ray spectrum during the low intensity and out of the low intensity state (left panel is with phase 0.31-0.33 and right panel is with phase 0.59-0.61 fitted with an absorbed power-law with gaussian model.

its partial eclipse. We looked at its orbital profile with *Swift*-BAT and XRT. The intensity dip during every orbital period observed in soft X-rays is similar to the intensity dip that has been seen in hard X-rays with *Swift*-BAT. The intensity falls off to $\sim 30\%$ during the eclipse.

Corbet et al. (2010) reported this significant modulation in the *Swift*-BAT lightcurve in every 4.24 days, and it was interpreted as an eclipse. This eclipse-like feature is not similar to the eclipse seen in other eclipsing HMXB systems, where the eclipse is usually broad, and the intensity during the eclipse falls to almost $\sim 1\%$ of that of the outside the eclipse. Aftab et al. (2019) studied the eclipse spectra of 11 HMXBs with the *XMM*-Newton and observed that the flux during the eclipse reduces by a factor of $\sim 8 - 237$. Also, Falanga et al. (2015) discussed 10 eclipsing HMXBs and presented the orbital intensity profile of the eclipsing HMXB sources like Cen X-3, 4U 1700-37, Vela X-1 etc. In Table 3.3 we have tabulated the average normalized count rate for inside and outside the eclipse phase observed in 8 of those HMXBs from *Swift*-BAT lightcurve. We did not include XTE J1855-026 and EXO 1722-363, as the eclipse count rate measurement of these sources have large relative error. To make Table 3.3, we took the most recent long-term orbital lightcurve from *Swift*-BAT site and folded with the orbital periods and epoch times reported in (Falanga et al. 2015). The eclipse profile of these HMXBs along with the ratio of the count rates, makes it clearly evident that the partial eclipse seen in IGR J16393-4643 is narrower and much shallower. This casts doubt if the partial eclipse

Table 3.3: Ratio of Swift-BAT normalized count rate of inside and outside eclipse phase of the HMXBs reported in (Falanga et al. 2015).

Name of the source	Normalized count rate (OOE ^a)	Normalized count rate (E ^b)	Ratio (OOE/E)
Vela X-1	1.3	2.7e-2	48.15
LMC X-4	1.6	3.3e-2	48.48
Cen X-3	1.3	2.5e-2	52.00
4U 1700-377	1.3	3.4e-2	38.24
4U 1538-522	1.2	6.9e-2	17.39
SMC X-1	1.2	6.5e-2	18.46
EXO 1722-363	1.2	4.1e-2	29.26
OA0 1657-415	1.4	8.7e-2	16.09

^a : Outside of eclipse phase excluding the eclipse ingress and egress

^b : Inside the eclipse phase

is really an eclipse or it is an X-ray dip taking place as a result of a periodic increase in absorption.

3.4.1 Variation in the spectral parameter

From the *Swift*-XRT observations, we extracted spectra in both inside and outside the low intensity phase (or the partial eclipse phase). We compared the spectral parameters in these phases.

3.4.1.1 Photon index

In other eclipsing X-ray binaries, the photon index is usually higher during the eclipse. It should be higher because during this time we only see the scattered emission from the wind. One such example is Cen X-3, where the correlation between the photon index

and the orbital phase can be clearly seen (Naik et al. 2011a). The spectral analysis of IGR J16393-4643 shows that the photon index during the eclipse is comparable to the same outside eclipse. If at all, the photon index is smaller during the partial eclipse. Although, the study by Aftab et al. (2019) found that the photon indices were lower during the eclipse for several HMXBs. Some HMXBs like SMC X-1, 4U 1700-377, even showed negative values of the photon indices during the eclipse.

3.4.1.2 The equivalent width of the Iron K alpha line

In high mass X-ray binaries, the companion star is either a super giant or a main sequence star. When it is a supergiant, the X-ray emitting compact object is deeply embedded in the dense wind of its companion. In these systems iron fluorescence lines are often present. They originate when X-ray photons interact with the absorbing matter in the stellar wind. Therefore, these lines are an indicator of the geometry and density of the matter around the compact object. The equivalent width of the fluorescent Fe line has a distinct correlation with the column density (Kallman et al. 2004). For isotropic distribution of cold matter this relation is linear and it can deviate from linearity for an anisotropic distribution. During the eclipse, X-ray spectra of HMXBs are known to exhibit a larger equivalent width of the iron $K\alpha$ line (Aftab et al. 2019) compared to the out of eclipse phase. It can be up to as much as 1.5 keV (Cen X-3: Nagase et al. 1992a, Naik et al. 2011a; Vela X-1: Nagase et al. 1994). In spite of the limited statistics of the data we were able to put an upper limit of ~ 350 eV to the equivalent width of the Fe- $K\alpha$ line present at 6.4 keV, which is considerably smaller than the iron line equivalent width of HMXBs in eclipse.

3.4.1.3 The absorption column density

If the partial eclipse is actually an X-ray dip resulting due to the increased absorption in the stellar wind, then an increased value of column density (N_H) is expected during the low intensity phase. From our spectral analysis, we do not observe any significant increase in the absorption column density during the low intensity state. It can not be confirmed whether the periodic decrease in intensity is a result of increased absorption or not.

In particular, this partial eclipse is very similar to the partial eclipse present in another HMXB 4U 0114+65 (a.k.a. 3A 0114+650), which was also interpreted previously

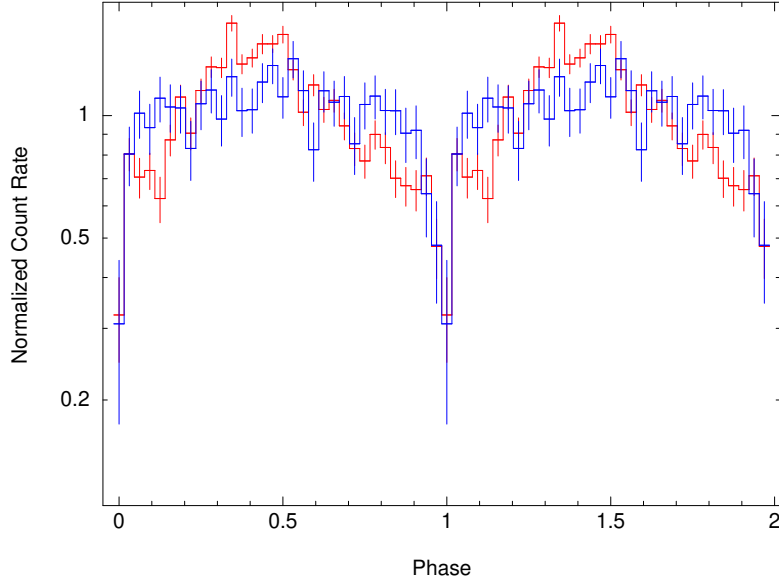


Figure 3.6: The *Swift*-BAT long term lightcurves of 4U 0114+65 (red) and IGR J16393-4643 (blue) folded with the period of 11.596 days and 4.237 days respectively.

as an eclipse in literature with an orbital period of ~ 11.59 days (Crampton et al. 1985, Grundstrom et al. 2007, Wen et al. 2006). In Figure 3.6, We have presented the *Swift*-BAT long-term lightcurve of 4U 0114+65 after folding it with 1001903.4 seconds (~ 11.596 days) with an epoch of MJD 53421.5. We can see the similarity in the intensity profile of IGR J16393-4643 and 4U 0114+65, which are plotted using *Swift*-BAT long-term observation of these two sources. However, spectral analysis of the low intensity feature of 4U 0114+65 with a *Suzaku* observation gave a moderate value of absorption column density and a low value of the equivalent width of Fe-K α line during the eclipse, which was in contrast with the high value of the equivalent width of Fe-K α line seen in other eclipsing HMXB spectra. These spectral characteristics along with detection of pulsations in low intensity state showed that this low intensity phase in 4U 0114+65 is not due to the eclipse, but it is an X-ray dip occurring due to the increased absorption in the stellar wind (Pradhan et al. 2015). Lower intensity during some narrow orbital phase range has also been reported in GX 301-2 (Islam and Paul 2014). The low intensity phase seen in IGR J16393-4643 is most likely due to absorption in the stellar corona. A grazing eclipse by the companion star occurs due to absorption caused by the stellar corona.

3.4.2 Inclination angle

To estimate the inclination angles for grazing eclipse caused by the stellar corona, we have followed the same method as in [Islam et al. \(2015\)](#). Using Kepler's law for a nearly circularized orbit, the orbital separation (a) between the companion star and the neutron star can be written in terms of their total mass ($M = M_c + M_{ns}$, where M_c is the mass of the companion and M_{ns} is the mass of the neutron star) and orbital period P_{orb} .

$$a^3 = \frac{G(M_c + M_{ns})P_{orb}^2}{4\pi^2} \quad (3.1)$$

Taking a standard value of $M_{ns} = 1.4M_\odot$ and $P_{orb} = 4.2$ days, a can be calculated for given value of M_c . The masses of the optical counterpart in the supergiant HMXBs like Vela X-1, Cen X-3 and OAO 1657-415 have been found to be in the range of $10 M_\odot - 30 M_\odot$ with their radii within the range of $10 R_\odot - 35 R_\odot$, in almost circular orbit ([Ash et al. 1999](#), [Mason et al. 2012](#), [Rappaport et al. 1983](#), [van Kerkwijk et al. 1995a](#)). Also, the main sequence B type stars in a circular binary orbit have masses within the range of $2 M_\odot - 16 M_\odot$ and radii in the range of $5 R_\odot - 25 R_\odot$ ([Habets and Heintze 1981](#)). The inclination angle of the binary can be calculated from the radius of the companion (Equation 5 of [Rubin et al. \(1996\)](#)).

$$\frac{R_c}{a} = \sqrt{\cos^2 i + \sin^2 i \sin^2 \Theta_E} \quad (3.2)$$

For a grazing eclipse $\Theta_E = 0$. Then equation 3.2 reduces to,

$$\frac{R_c}{a} = \sqrt{\cos^2 i} \quad (3.3)$$

Also, the upper limit on the radius of the companion is it's Roche lobe radius $R_{c,roche}$. An analytical approximation (accurate $<1\%$) of $R_{c,roche}$ ([Bowers and Deeming 1984](#), [Hill et al. 2005](#)) is,

$$R_{c,roche} = a(0.38 + 0.2 \log(\frac{M_c}{M_{ns}})) \quad (3.4)$$

Equation 3.4 puts a further constraint on the range of the inclination angle. The optical counterpart of the X-ray pulsar in IGR J16393-4643 is either a supergiant O9 star with $R \sim 20R_\odot$ or a main sequence B star with $R \sim 10R_\odot$ ([Bodaghee et al. 2012](#)). We used the above equations and took an eclipse semi-angle of zero for the nearly circular orbit. In Fig. 3.7, the left and right panels show the inclination angles estimated for the given

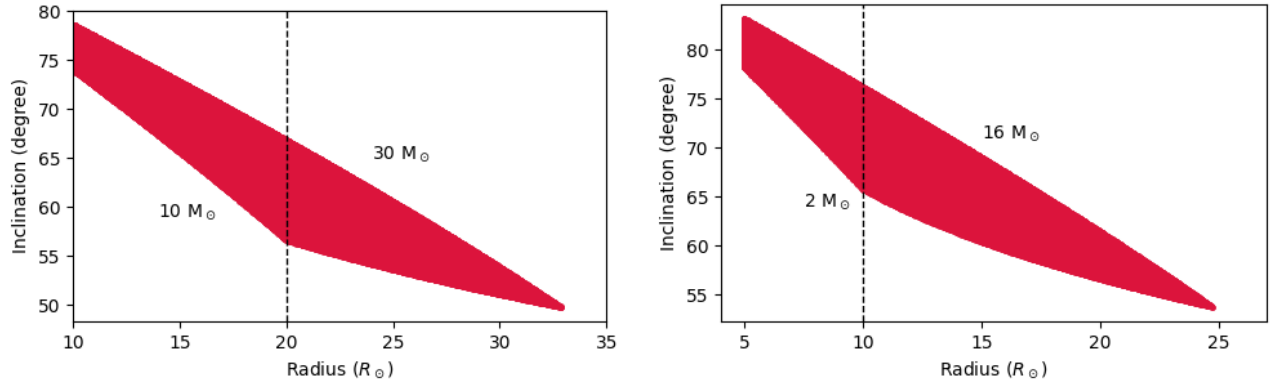


Figure 3.7: The plots shows the inclination angles estimated for the given mass ranges of the O9 supergiant companion (left panel) and the main sequence B star (right panel) as a function of its radius. The dotted line in the left panel represents the $20 R_{\odot}$ radius for a O9 supergiant companion and the dotted line in the right panel represents the $10 R_{\odot}$ radius for a B type companion.

mass ranges of the O9 supergiant companion and the main sequence B star respectively as a function of its radius. The estimated inclination angle is $65^{\circ} - 77^{\circ}$ for a main sequence B star with $10 R_{\odot}$ and the range of $56^{\circ} - 68^{\circ}$ for a supergiant O9 star with $20 R_{\odot}$. The dotted vertical lines in Figure 3.7 represent the $10 R_{\odot}$ radius for the main sequence B star (right panel) and $20 R_{\odot}$ radius for the supergiant O9 star (left panel).

Chapter 4

Broadband X-ray characteristics of the transient pulsar GRO J2058+42



Based on:

Broadband X-ray characteristics of the transient pulsar GRO J2058+42

Kabiraj S. & Paul B., MNRAS **497**, 1059, 2020

4.1 Introduction

In X-ray binary pulsars (XBPs) the neutron star accretes matter from its companion and the in-falling matter after crossing the Alfvén radius (radius inside which the magnetic stress influences the flow in the accretion disk) is channeled to the magnetic field lines and finally reaches the poles of the neutron star. The gravitational potential energy of the accreted matter powers X-ray radiation and the X-ray luminosity is governed by the mass accretion rate. The angular momentum of the accreted matter is added to the angular momentum of the neutron star, causing it to rotate faster. Therefore the spin-up rate and X-ray luminosity are expected to be correlated, and this relation depends on various factors like mass, radius, magnetic field of the neutron star, mode of accretion, etc. According to the standard model of accretion, $\dot{\nu} \propto L_X^{6/7}$ (Ghosh and Lamb 1979b), where $\dot{\nu}$ is the spin frequency change rate ($= -\frac{\dot{P}}{P^2}$, P is the spin period).

The Be X-ray binary systems usually harbour a neutron star as the compact object. They have a wide eccentric orbit and a gaseous circumstellar geometrically thin disk in Keplerian motion around the equator of the Be type companion star. The neutron star accretes matter from its companion's slow, dense equatorial stellar outflow. Such systems are often dormant or show weak outbursts in every orbit. However, sometimes the outflow from the companion star is enhanced and when the neutron star crosses the disk, giant outbursts for weeks may occur. An accretion disk is often formed during these giant outbursts. As a result, rapid spin up episode along with large increase in the X-ray luminosity is observed. This makes the Be X-ray binaries excellent candidates to establish a quantitative comparison of theoretical predictions with minute observations. The luminosity and the spin frequency derivatives during large outbursts with $L_X \geq 1 \times 10^{37} \text{ ergs}^{-1}$ were found to be related as $\dot{\nu} \propto L_X^\alpha$ ($\alpha \simeq 1$) (Bildsten et al. 1997, Parmar et al. 1989b, Sugizaki et al. 2017). This proportionality relation validates the expected theoretical XBP spin up models of magnetospheric interaction with accretion disk. The giant outbursts also provide an excellent opportunity to characterize the broadband pulsation and spectral properties, especially with observatories like *NuSTAR*. Detection of a Cyclotron Resonance Scattering Feature (CRSF) in the X-ray spectrum helps to determine the magnetic field strength of the neutron star unambiguously. In this chapter, we study one such Be XBP called GRO J2058+42.

4.2 GRO J2058+42

GRO J2058+42 (a.k.a. CXOU J205847.5+414637) is a transient high mass X-ray binary pulsar. It is a slowly rotating pulsar with a spin period of ~ 198 seconds. It was discovered during a giant outburst with BATSE onboard the Compton Gamma Ray Observatory in September, 1995 (Wilson et al. 1998; 1995). The outburst lasted for about 46 days until October 30, 1995, and spun up from a period of 198 seconds to 196 seconds during the outburst with the pulsed flux peaking at 140 mCrab (20-50 keV). The initial giant outburst was followed by a series of normal outbursts with pulsed flux peaking at 15-20 mCrab (20-50 keV) occurring at ~ 110 days intervals. BATSE also detected shorter and weaker outbursts with peak-pulsed fluxes of ~ 8 mCrab in between the 15-20 mCrab outbursts. These normal and weaker outbursts were similar in intensity when observed with *RXTE*-ASM indicating an orbital cycle of 55 days (Corbet et al. 1997). An alternate explanation was that the normal and weaker outbursts are related to two different accretion mechanism during the periastron and apastron passage and the orbital cycle is of ~ 110 days (Bildsten et al. 1997, Wilson et al. 1998). The occurrence of periodic outbursts after a giant outburst indicates the companion of GRO J2058+42 is likely to be a Be star.

With the optical and X-ray observations carried out during quiescence, GRO J2058+42 was proven to be a Be X-ray binary (Reig et al. 2005, Wilson et al. 2005). Reig et al. (2005) identified the companion as an O9.5-B0 IV-Ve star with $V \simeq 14.9$ mag and $R \simeq 14.2$ mag at a distance of $\sim 9 \pm 1$ kpc, which is in agreement with the distance range of 7-16 kpc estimated from accretion torque during the first giant outburst (Wilson et al. 1998). In May, 2008 another outburst activity from GRO J2058+42 was detected with *Swift*-BAT hard X-ray transient monitor in 15-50 keV energy band. The BAT flux was similar to the normal outbursts and it was not accompanied by any giant outburst.

Recently, a second giant outburst from the source has been observed with *Swift*-BAT and *Fermi*-GBM in the mid March of 2019. *Swift*-BAT triggered at this source on March 22 (Barthelmy et al. 2019) and on March 27 (Kennea et al. 2019, Lien and Page 2019) reporting a renewed activity. Malacaria et al. (2019) reported the GBM detection of the source, going through a bright outburst starting on March 14, 2019 reaching to a peak luminosity similar to the previous giant outburst in 1995 observed by BATSE. It was detected with GBM for ~ 52 days. In this period, *NuSTAR* carried out two target of opportunity (TOO) observation on March 25 and April 11, 2019.

Among the bright transient Be X-ray binaries, GRO J2058+42 is a relatively poorly studied source. Using the GBM and *NuSTAR* observation of the recent second giant outburst of this source, we have looked at its broadband X-ray spectrum and searched for the CRSF feature, if any, and studied the spin-up rate and pulsed X-ray luminosity to estimate the magnetic field (B field) strength of the pulsar. The spin-up characteristics was previously studied by [Wilson et al. \(1998\)](#) from its first giant outburst observation but the distance of the binary was still unknown. So it was not possible to estimate the B field earlier. We have also analyzed the pulse profiles in broad energy band from two *NuSTAR* observations to look at their pulse shapes in order to find any change in the accretion mode.

4.3 Observation

4.3.1 *NuSTAR* Observation

Two target of opportunity (TOO) observation of GRO J2058+42 during the outburst were made by *NuSTAR* on Mar 25, 2019 for 20.4 ks (ObsID 90501313002, $T_{start} = 58567.30$ MJD) and on April 11, 2019 for 38.6 ks (ObsID 90501313004, $T_{start} = 58584.01$ MJD). Also after the outburst, there was another *NuSTAR* observation on August 28, 2019 for 38.5 ks (ObsID 90501336002, $T_{start} = 58723.53$) when the source luminosity had decreased considerably. To process and filter the preliminary *NuSTAR* data we have used HEASOFT version 6.25 in which NuSTARDAS pipeline version 1.8.0 is installed as a sub-package. First using standard NuSTARDAS pipeline, we extracted the clean event files. Then using DS9 version 8.0.1, 120" radius region centering the source was selected. It was used to extract the source photons in order to create the source lightcurve and spectra. Another circular region of 120" away from the source in the FoV was selected to create background lightcurve and spectra. The response matrix files and ancillary files were generated using CALDB version 1.0.2.

4.3.2 *Fermi*-GBM Observation

In this work, we have used the pulse frequency and 12-50 keV pulsed flux measurements with the *Fermi*-GBM ([Camero-Arranz et al. 2009](#), [Finger et al. 2009](#)). The measurement technique is discussed thoroughly in [Malacaria et al. \(2020\)](#). GRO J2058+42 was detected

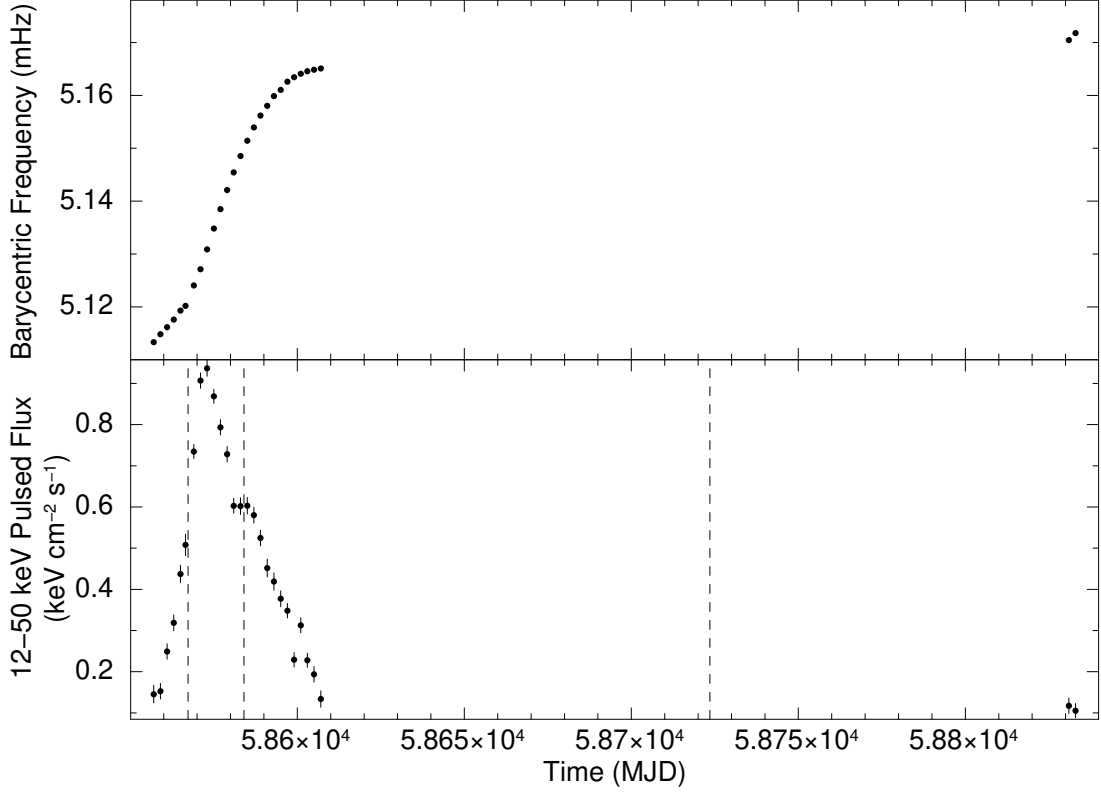


Figure 4.1: Top: barycentric spin-frequency history measured with GBM. The typical error associated with each pulse frequency measurement is of the order of 10^{-4} mHz, which is too small to be visible in this plot. Bottom: pulsed flux measured with GBM in 12-50 keV energy band. The three vertical lines represent the times of the three *NuSTAR* observations.

with GBM from Mar 14, 2019 to May 4, 2019 for a period of about 50 days during its most recent outburst. Figure 4.1 represents the barycentric pulse frequency and pulsed flux profile with time ¹. The error bars in figure 4.1 stands for the associated uncertainties in these measurements due to statistical limitations and time intervals. The errors in pulse frequency measurements are about four orders of magnitude smaller than the pulse frequency values and are smaller than the symbols used in the figure 4.1.

4.4 Broadband X-ray characteristics

4.4.1 Timing Analysis

We created source and background lightcurves from the two *NuSTAR* observations with a bin time of 1 second for both FPMA and FPMB. For each observation, we added the

¹<https://gammaray.msfc.nasa.gov/gbm/science/pulsars/lightcurves/groj2058.html>

lightcurves from FPMA and FPMB after background subtraction. We determined the pulse periods (P_{spin}) during the three observations by the method of epoch folding and chi-square maximization using **efsearch** v1.1 (XRONOS v5.22). The pulse periods at reference epoch MJD 58567.30 and MJD 58584.01 were found to be 195.25 ± 0.02 seconds and 194.14 ± 0.02 seconds respectively. These two periods are consistent with the pulse period evolution recorded with *Fermi*-GBM during the outburst. The pulse period at reference epoch MJD 58723.53 was found to be 193.37 seconds.

Then, following the same method we generated 6 background-subtracted lightcurves in 3-6 keV, 6-12 keV, 12-20 keV, 20-30 keV, 30-50 keV and 50-78 keV energy band for each of these three observations and folded them with their respective pulse periods (Figure 4.2, 4.3, 4.4). Then we aligned and overlaid these energy resolved pulse profiles from three observation on top of each other (Figure 4.5) to observe any change in the pulse shapes. For the observations during the outburst, the profiles do not look significantly different from each other. The pulse shapes has strong energy dependence. In the range of 3-12 keV the pulse profile has at least four peaks and the two main peaks merged into a single peak above 30 keV. The energy dependence of the pulse shapes are consistent with the previous *RXTE*-PCA observation in November 28, 1996 (Wilson et al. 1998). The pulse profiles from the third observation (when the pulsar is in the low luminosity state) have different shapes as compared to the first two. They have three peaks in the lowest energy band (3-6 keV). As we approach the higher energy bands, it becomes double peaked above 20 keV. The error bars are large above 50 keV.

4.4.2 Spectral Analysis

We have extracted the source and background pulse phase-averaged spectra and response files from the three *NuSTAR* observations using **nuproducts** the command in the 3-78 keV energy band. We have fitted FPMA and FPMB spectra with the same spectral model but allowed the relative normalization of the detectors to vary. First we tried various continuum models used for HMXB pulsars like cutoff power-law, high energy cutoff power-law (HIGHECUT; White et al. 1983), NewHcut (a third order polynomial function with continuous derivatives; Burderi et al. 2000), Fermi Dirac cutoff power-law (FD-CUT; Tanaka 1986), Thermal Comptonization model (CompTT; Titarchuk 1994) to fit the spectra.

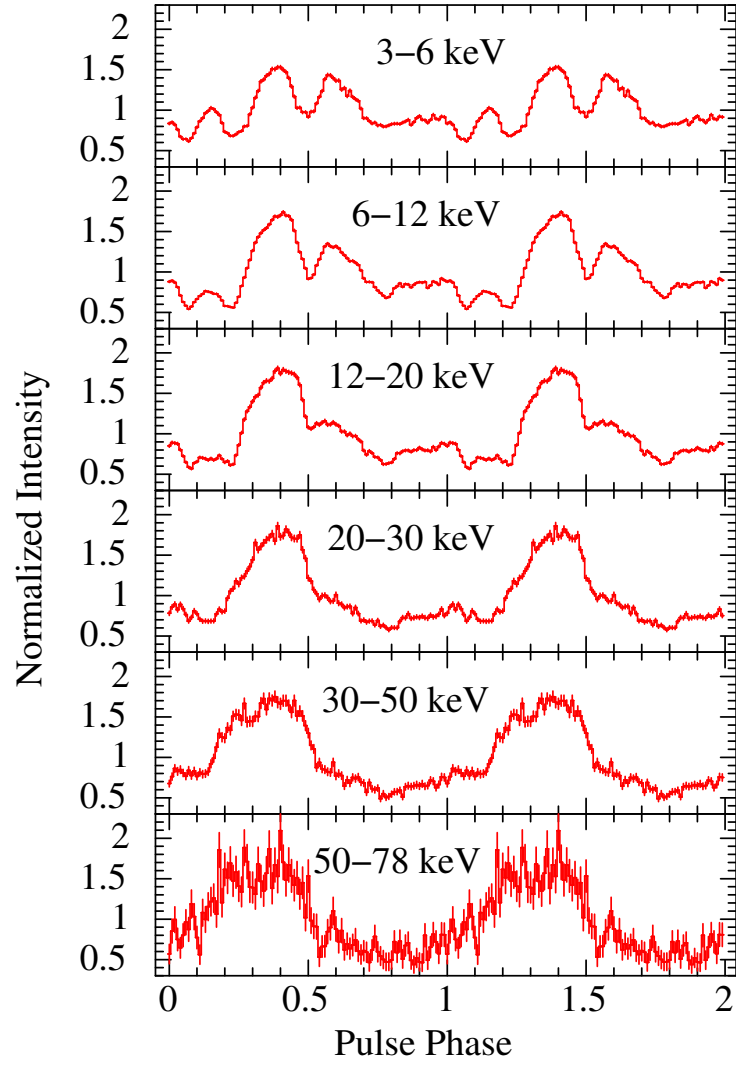


Figure 4.2: Energy resolved pulsed profiles from *NuSTAR* observation with Obs ID 90501313002 (MJD 58567.30) folded with $P_{spin} = 195.25$ seconds.

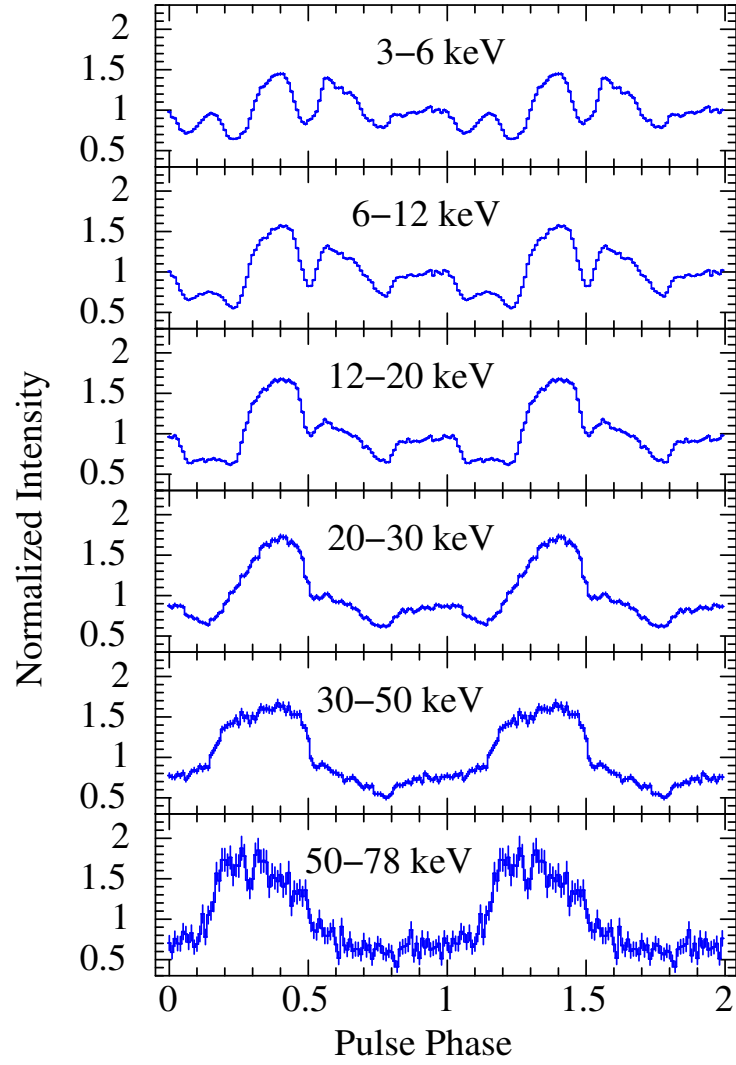


Figure 4.3: Energy resolved pulsed profiles from *NuSTAR* observations with Obs ID 90501313004 (MJD 58584.01) folded with $P_{spin} = 194.14$ seconds.

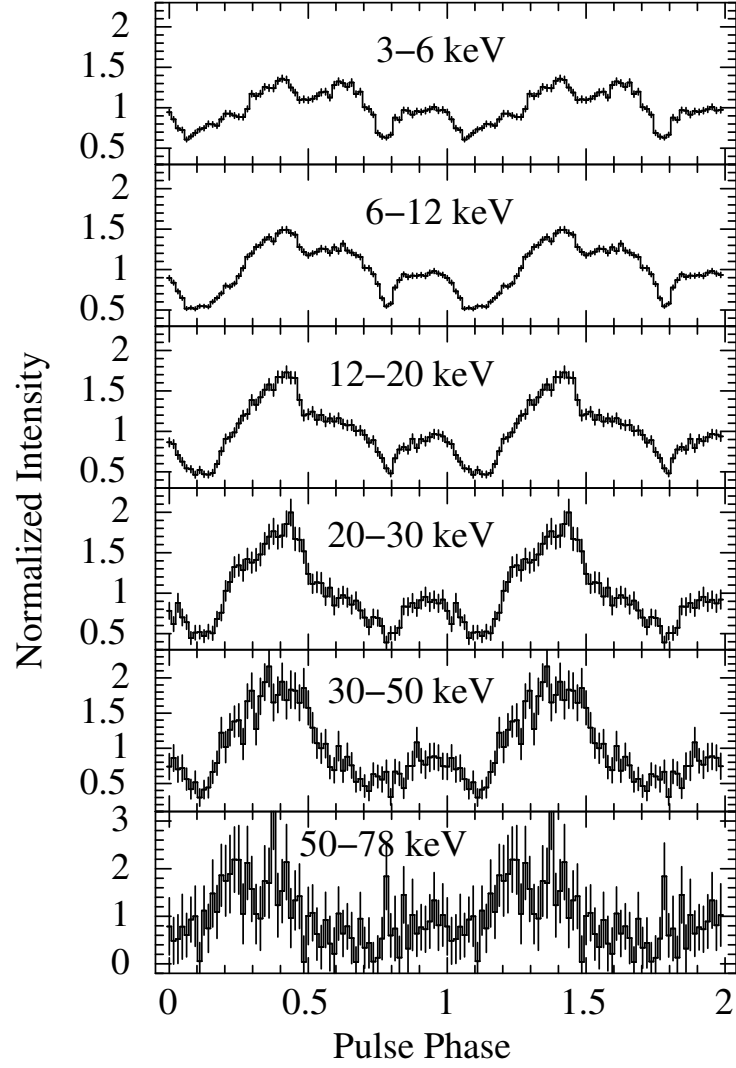


Figure 4.4: Energy resolved pulsed profiles from *NuSTAR* observations with Obs ID 90501336002 (MJD 58723.53) folded with $P_{spin} = 193.37$ seconds.

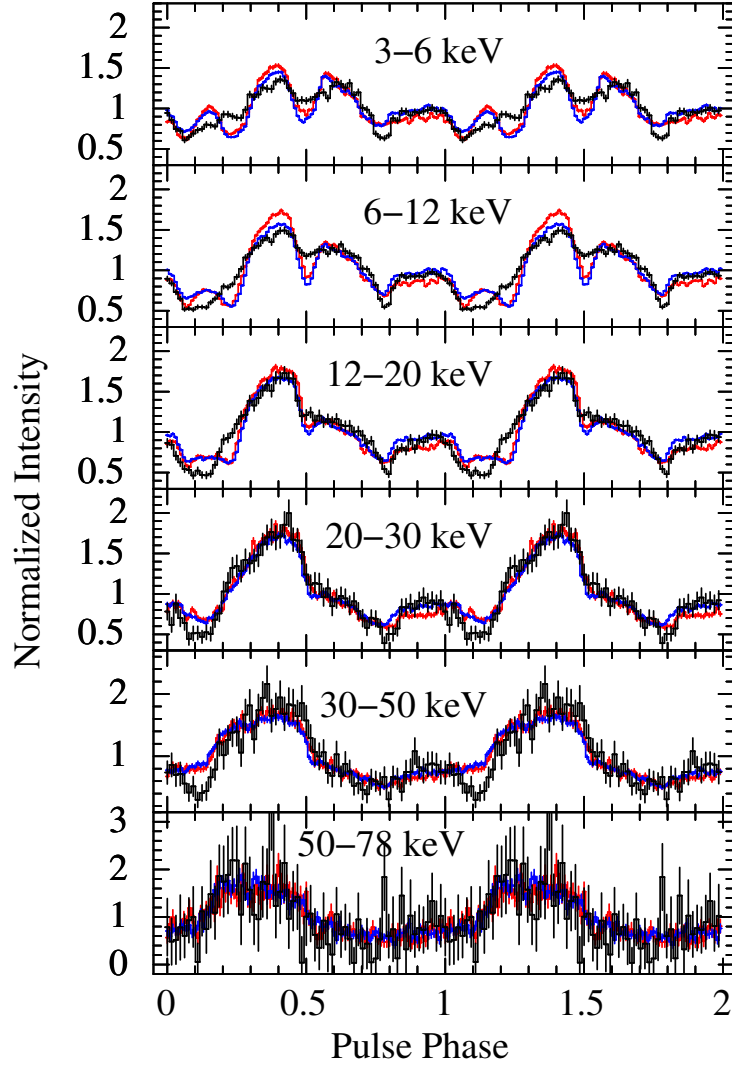


Figure 4.5: Energy resolved pulsed profiles from all the three *NuSTAR* observations with Obs ID 90501313002 (red), Obs ID 90501313004 (blue) and Obs ID 90501336002 (magenta) folded with $P_{spin} = 195.25$ seconds, 194.14 seconds and 193.37 seconds respectively.

4.4.2.1 During the outburst

The HIGHECUT, NewHcut and Fermi Dirac cutoff power-law spectral models described the source spectra during the outburst well with physically acceptable parameters. For all these three cases, Fe-K α line at 6.4 keV is present in both the spectra with an equivalent width of about 125 eV. However, when fitted with the HIGHECUT and NewHcut cut-off power law model, the spectra required an additional absorption component (gaussian absorption line ¹) which indicated possible presence of a CRSF feature at around 25 keV. But in all such cases, the cyclotron energy is almost same as the cut-off energy, which implicate the CRSF feature to be an artifact. FD-CUT model does not show presence of any negative residual. and provided the best fit to the spectra with minimum χ^2 . While fitting the spectra, we considered a component *phabs* to account for the absorption in the interstellar medium along our line of site. But the value of the equivalent hydrogen column density (N_H) came out to be smaller than the galactic N_H and it could not be constrained. Therefore, we have included neutral hydrogen absorption fixed to the Galactic value of $N_H = 6.19 \times 10^{21} \text{ cm}^{-2}$ calculated using HEASARC tools ². The spectral parameters for these models are tabulated in Table 4.1 and 4.2 .

Figure 4.6 and 4.7 shows the overlaid FPMA and FPMB spectra from the two observations by *NuSTAR*. They are fitted with *constant*phabs(powerlaw*fdcut + gaussian)* where the *constant* is to allow for a relative normalization for FPMA and FPMB. It is kept as 1 for FPMA and varied for FPMB. *phabs* N_H is to account for the galactic value. A *gaussian* component is used to fit the Iron line at 6.4 keV and *powerlaw*fdcut* is used to fit the continuum. The second, third and fourth panels in the Figure 4.6 and 4.7 are the χ when the cutoff energy components are FDCUT, HIGHECUT and NEWHCUT respectively. These fits do not include a gaussian absorption line in the model.

We also fitted the broadband spectra with the best fit model and included an extra absorption component (*gabs*) at a particular energy and repeated this for every energy in the *NuSTAR* band. We determined the upper limit on the optical depth each time and presented it in Figure 4.9. To determine the upper limit of any absorption line in the spectrum, the width of the line was taken to be 1 keV and 2 keV in the energy bands of 4-10 keV and 11-78 keV respectively. Since the spectrum is only up to 80 keV and statistics is poor at higher energy, it is not well constrained beyond 60 keV.

¹<https://heasarc.gsfc.nasa.gov/xanadu/xspec/manual/node240.html>

²<https://heasarc.gsfc.nasa.gov/cgi-bin/Tools/w3nh/w3nh.pl>

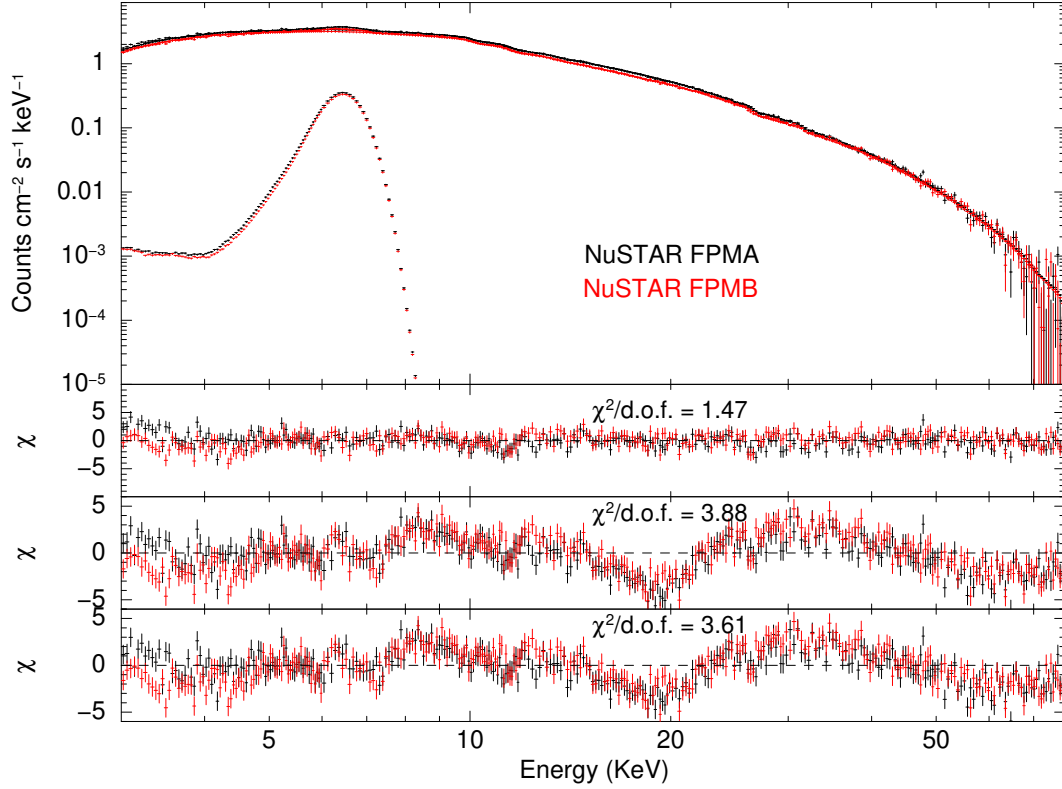


Figure 4.6: The first panel shows the overlaid X-ray spectra of FPMA and FPMB from NuSTAR observation with ObsID 90501313002 modelled with absorbed power law and FDCUT along with an emission line. The second, third and fourth panels represent the (data-model)/error (i.e. χ) when the cut-off energy component is FDCUT, HIGHECUT and NEWHCUT respectively.

4.4.2.2 During the low luminosity state

The spectra from the third *NuSTAR* observation during the quiescent state was best fitted with an absorbed power law with HIGHECUT. The fit was almost equally good with NewHcut. We did not see the presence of any emission or absorption feature in both the cases. We got an higher value of the absorption column density (N_H) in this state. The photon index increased almost by a factor of ~ 2 and the flux dropped by a factor of ~ 10 . The fit was poorer with FDCUT. With FDCUT, the fitting parameters could not be constrained within a physically acceptable range. We have tabulated the spectral parameters for these models in Table 4.3.

Figure 4.8 shows the overlaid FPMA and FPMB spectra from the *NuSTAR* observation during the low luminosity state. They are fitted with *constant*phabs(powerlaw*highecut)*.

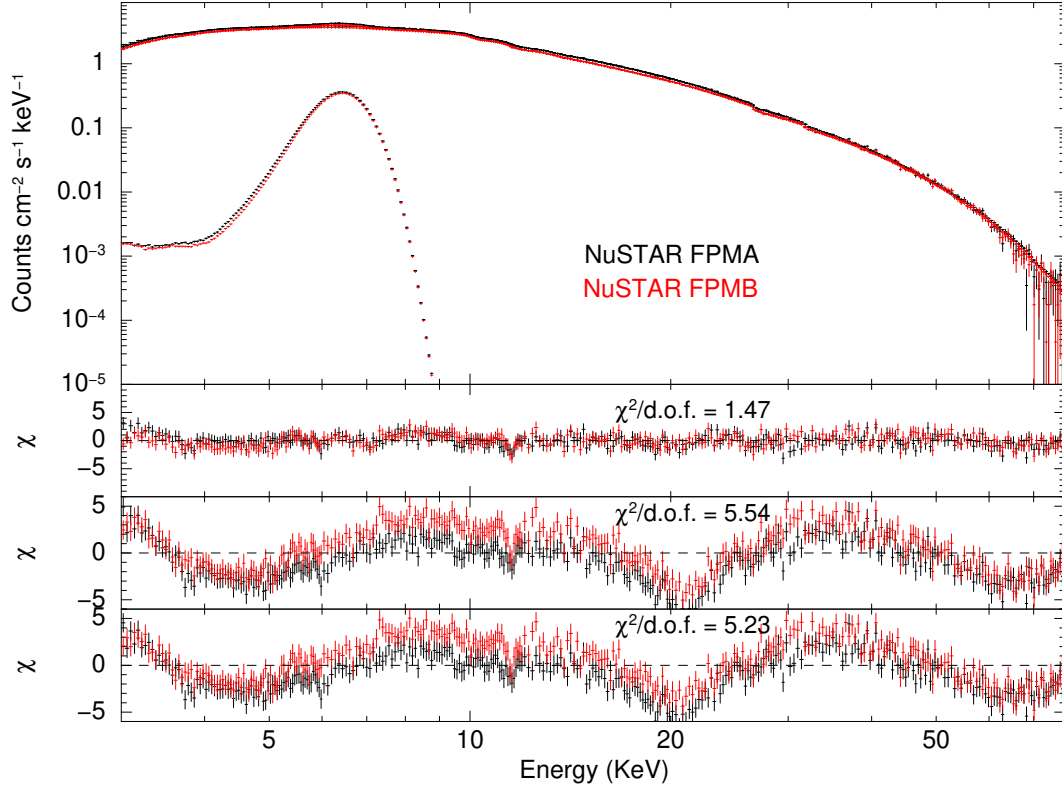


Figure 4.7: The first panel shows the overlaid X-ray spectra of FPMA and FPMB from NuSTAR observation with ObsID 90501313004 modelled with absorbed power law and FDCUT along with an emission line. The second, third and fourth panels represent the (data-model)/error (i.e. χ) when the cut-off energy component is FDCUT, HIGHECUT and NEWHCUT respectively.

The second, third and fourth panels in the Figure 4.7 are the χ when the cutoff energy components are HIGHECUT and NEWHCUT respectively.

4.5 Accretion torque and estimation of the magnetic field strength

4.5.1 Spin change rate and luminosity measurement

The pulsed frequencies (ν_s) measured with *Fermi*-GBM clearly show a rapid spin-up of the pulsar during the second giant outburst (Figure 4.1). The total increase in frequency during the second outburst is about 0.05 mHz which is almost similar to 0.06 mHz during the first giant outburst. In total there were 26 ν_s measurements taken roughly in every 2

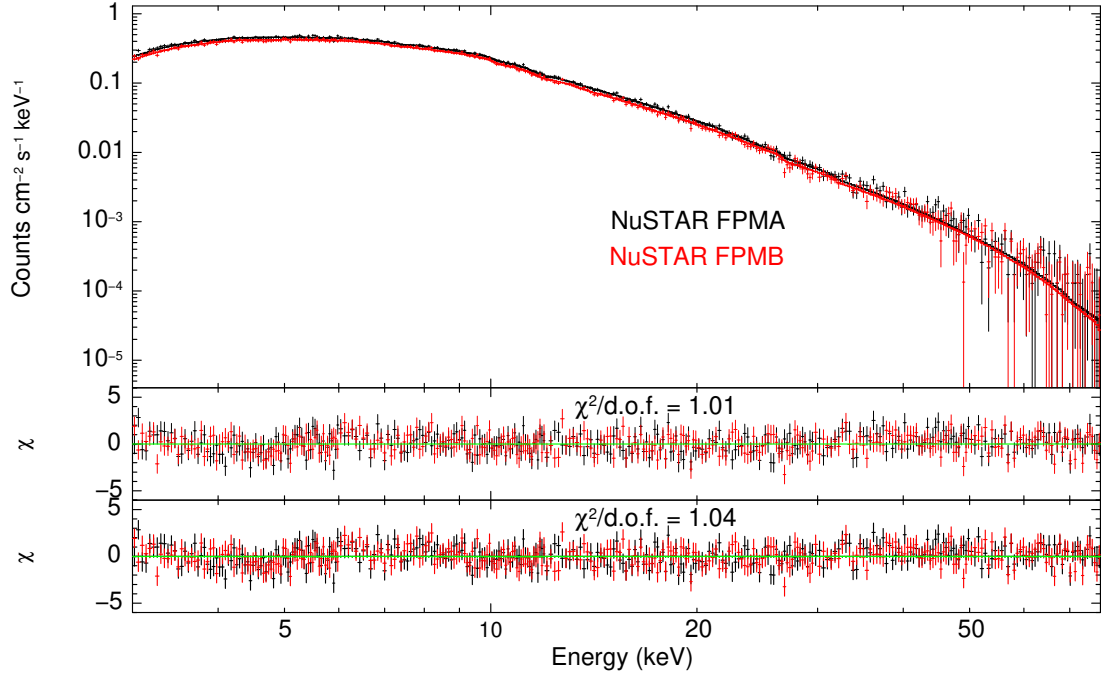


Figure 4.8: The first panel shows the overlaid X-ray spectra of FPMA and FPMB from NuSTAR observation with ObsID 90501336002 modelled with absorbed power law and HIGHECUT. The second, third and fourth panels represent the (data-model)/error (i.e. χ) when the cut-off energy component is HIGHECUT and NEWHCUT respectively.

day intervals. We took first three consecutive ν_s and fitted them w.r.t. time with a linear function. While fitting, we took into account the uncertainties associated with the pulse frequency measurements and because the integration time for each pulsation value is 2 days, the uncertainties associated with time is taken to be 1 day. We achieved the best fit using χ^2 minimization technique and determined the spin up rate from the slope of the linear function for 6 days interval, as each pulse frequency measurement is taken in every two days interval. We estimated the 1σ statistical error on spin-up rate ($\dot{\nu}_s$). We repeated this process for the next three ν_s and so on. Therefore, we had 8 spin-up rates ($\dot{\nu}_s$) from 24 spin frequencies (ν_s) and we did not use the last 2 ν_s .

From the *NuSTAR* data we estimated total flux (3-78 keV) at the time of the *NuSTAR* observations. We determined a conversion factor by which a pulsed flux measurement is related to the total flux from the *NuSTAR* observation at a particular time. Then we multiplied all the 26 pulsed flux detected with *Fermi*-GBM with this conversion factor to convert them into the total flux. To make a simultaneous sampling of the luminosity and spin-up rate measurements, we have taken the average value of the total flux over the

Table 4.1: Spectral parameters with quoted errors for 90% confidence limits.

parameter	ObsID 90501313002		
	FD-CUT	HIGHECUT	NewHcut
N_H^a (fixed)	6.19×10^{-1}	6.19×10^{-1}	6.19×10^{-1}
Flux ^b	$3.67^{+0.01}_{-0.01}$	$3.68^{+0.01}_{-0.01}$	$3.68^{+0.02}_{-0.01}$
continuum			
Photon Index (Γ)	$0.88^{+0.01}_{-0.01}$	$0.95^{+0.01}_{-0.01}$	$0.95^{+0.01}_{-0.01}$
power-law normalization ^c	$6.27^{+0.05}_{-0.05}$	$6.11^{+0.08}_{-0.08}$	$6.11^{+0.08}_{-0.09}$
E_{cut} (keV)	$27.72^{+0.58}_{-0.60}$	$24.31^{+0.34}_{-0.36}$	$24.99^{+1.14}_{-0.79}$
E_{fold} (keV)	$11.76^{+0.20}_{-0.20}$	$14.83^{+0.28}_{-0.27}$	$14.33^{+0.49}_{-0.64}$
emission lines			
Fe-K α line energy (keV)	$6.46^{+0.03}_{-0.03}$	$6.46^{+0.03}_{-0.03}$	$6.46^{+0.03}_{-0.03}$
$\sigma_{K\alpha}$ (keV)	$0.36^{+0.04}_{-0.04}$	$0.42^{+0.05}_{-0.05}$	$0.42^{+0.05}_{-0.04}$
Equivalent width (eV)	114^{+8}_{-10}	136^{+9}_{-8}	133^{+9}_{-9}
E_{gabs} (keV)	—	$24.41^{+0.43}_{-0.43}$	$25.12^{+1.25}_{-0.89}$
σ_{gabs} (keV)	—	$5.62^{+0.32}_{-0.30}$	$6.14^{+0.48}_{-0.40}$
$strenth_{gabs}$	—	$3.67^{+0.38}_{-0.34}$	$4.23^{+0.90}_{-0.58}$
statistic			
$\chi^2/d.o.f.$	1.47^d	1.58^e (3.88 ^d)	1.56^e (3.61 ^d)
relative normalization			
FPMB	$1.023^{+0.003}_{-0.003}$	$1.023^{+0.003}_{-0.003}$	$1.023^{+0.003}_{-0.003}$

Table 4.2: Spectral parameters with quoted errors for 90% confidence limits.

parameter	ObsID 90501313004		
	FD-CUT	HIGHECUT	NewHcut
N_H^a (fixed)	6.19×10^{-1}	6.19×10^{-1}	6.19×10^{-1}
Flux ^b	$4.42^{+0.01}_{-0.01}$	$4.43^{+0.01}_{-0.01}$	$4.42^{+0.01}_{-0.01}$
continuum			
Photon Index (Γ)	$0.83^{+0.01}_{-0.01}$	$0.89^{+0.01}_{-0.01}$	$0.86^{+0.01}_{-0.01}$
power-law normalization ^c	$7.04^{+0.05}_{-0.05}$	$6.47^{+0.08}_{-0.09}$	$6.40^{+0.09}_{-0.09}$
E_{cut} (keV)	$24.79^{+0.58}_{-0.60}$	$24.26^{+0.35}_{-0.34}$	$24.47^{+0.52}_{-0.50}$
E_{fold} (keV)	$12.71^{+0.16}_{-0.16}$	$14.47^{+0.23}_{-0.23}$	$14.17^{+0.28}_{-0.28}$
emission lines			
Fe-K α line energy (keV)	$6.45^{+0.04}_{-0.04}$	$6.43^{+0.04}_{-0.04}$	$6.43^{+0.04}_{-0.04}$
$\sigma_{K\alpha}$ (keV)	$0.48^{+0.06}_{-0.05}$	$0.54^{+0.06}_{-0.06}$	$0.52^{+0.06}_{-0.06}$
Equivalent width (eV)	129^{+10}_{-11}	146^{+10}_{-10}	140^{+9}_{-9}
E_{gabs} (keV)	—	$24.48^{+0.40}_{-0.39}$	$24.71^{+0.57}_{-0.55}$
σ_{gabs} (keV)	—	$6.50^{+0.40}_{-0.39}$	$6.95^{+0.31}_{-0.30}$
$strenth_{gabs}$	—	$5.25^{+0.46}_{-0.42}$	$5.83^{+0.58}_{-0.53}$
statistic			
$\chi^2/d.o.f.$	1.13^d	1.19^e (5.54 ^d)	1.14^e (5.23 ^d)
relative normalization			
FPMB	$1.028^{+0.002}_{-0.002}$	$1.028^{+0.002}_{-0.002}$	$1.028^{+0.002}_{-0.002}$

Table 4.3: Spectral parameters with quoted errors for 90% confidence limits.

parameter	ObsID 90501336002	
	HIGHECUT	NewHcut
N_H^a	$2.78^{+0.33}_{-0.37}$	$2.79^{+0.33}_{-0.34}$
Flux ^b	$0.30^{+0.01}_{-0.01}$	$0.30^{+0.01}_{-0.01}$
continuum		
Photon Index (Γ)	$1.51^{+0.03}_{-0.04}$	$1.50^{+0.04}_{-0.04}$
power-law normalization ^c	$2.60^{+0.01}_{-0.01}$	$2.58^{+0.01}_{-0.01}$
E_{cut} (keV)	$7.71^{+0.36}_{-0.48}$	$7.63^{+0.50y}_{-0.58}$
E_{fold} (keV)	$30.76^{+0.36}_{-0.48}$	$30.03^{+2.14}_{-1.92}$
emission lines		
Fe-K α line energy (keV)	—	—
$\sigma_{K\alpha}$ (keV)	—	—
Equivalent width (eV)	—	—
E_{gabs} (keV)	—	—
σ_{gabs} (keV)	—	—
$strenth_{gabs}$	—	—
statistic		
$\chi^2/d.o.f.$	1.01^d	1.04^d
relative normalization		
FPMB	$1.004^{+0.006}_{-0.006}$	$1.004^{+0.006}_{-0.006}$

- a* : XSPEC normalization , units of 10^{22} atoms cm^{-2}
- b* : XSPEC normalization , units of 10^{-9} ergs $\text{cm}^{-2} \text{sec}^{-1}$ & in the 3-78 keV range
- c* : XSPEC normalization , units of 10^{-2} photons $\text{cm}^{-2} \text{s}^{-1}$ at 1 keV
- d* : without gabs
- e* : with gabs
- f* : Large error with ambiguous fit values

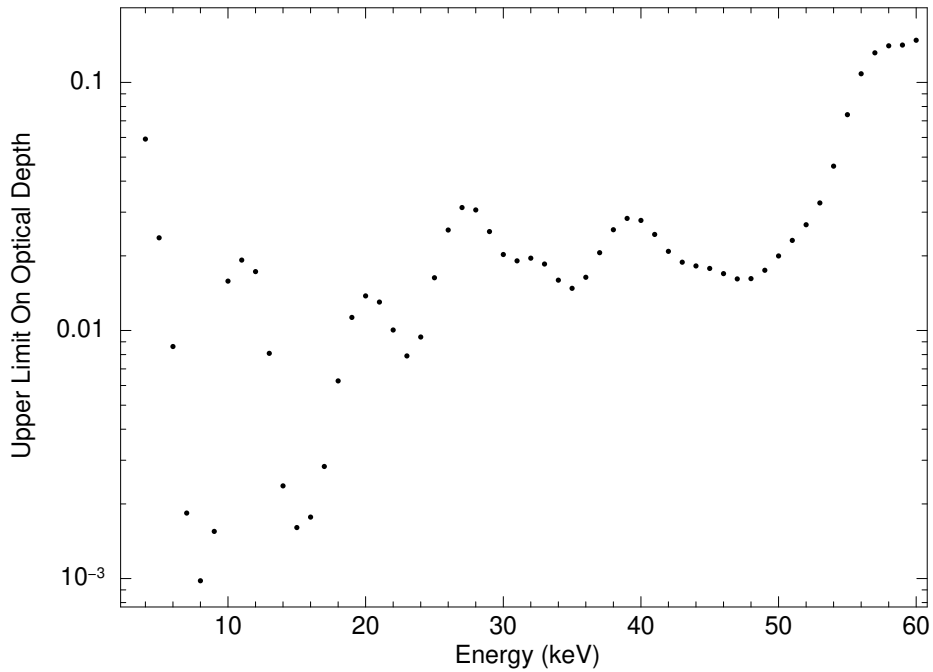


Figure 4.9: The upper limit on optical depth of a gaussian absorption line as a function of energy. This figure was generated from the *NuSTAR* observation with the ObsID 90501313002.

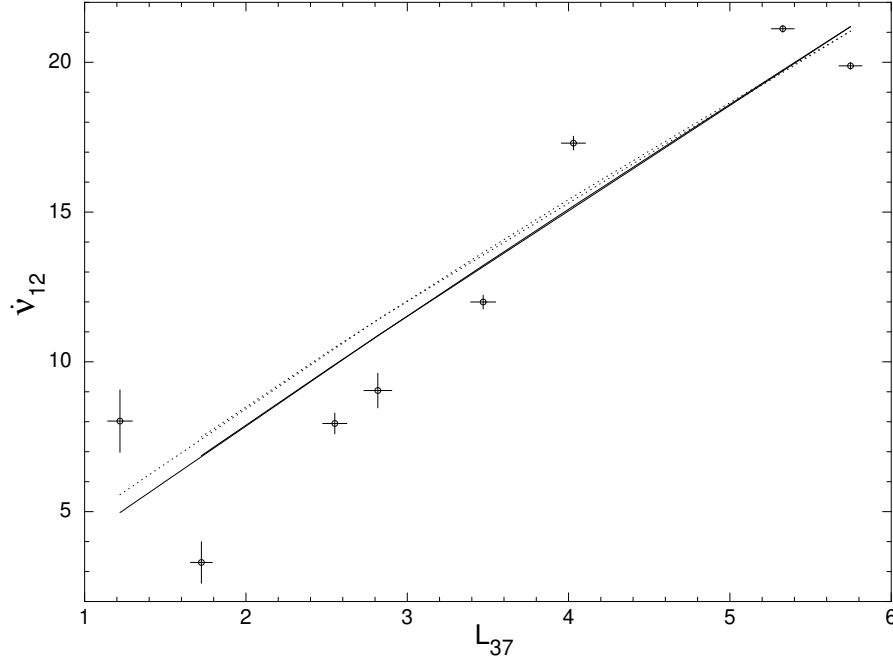


Figure 4.10: Plot of spin change rate and luminosity. The dotted line has power-law index of $6/7$ and the solid line represents the best power-law fit of the data points which gives a power-law index of 0.93 .

same 6 days intervals which are used to determine a particular $\dot{\nu}_s$. We then had 8 spin-up rates and total flux measurements during the outburst. We multiplied the uncertainties on the pulsed flux values by the same conversion factor and estimated the propagated error on the average value of the total fluxes. Then using the distance estimation of 9 kpc (Reig et al. 2005), we have calculated the luminosity, along with its uncertainty. A plot of the spin-up rate against bolometric luminosity is shown in Figure 4.10. We have made a power law fitting of $\dot{\nu}_s$ and luminosity to check the relation between them.

4.5.2 Spin change rate and luminosity relationship

The spin-up rate ($\dot{\nu}_s$) and the luminosity L are known to be correlated in transient X-ray pulsars (Sugizaki et al. 2017). The disk-magnetosphere interaction model of a rotating neutron star which is spinning up by accreting mass from its companion via a Keplerian disk, with a moment of inertia I , holds the following relation approximated to the non-relativistic limit (Ghosh and Lamb 1979b, Sugizaki et al. 2017).

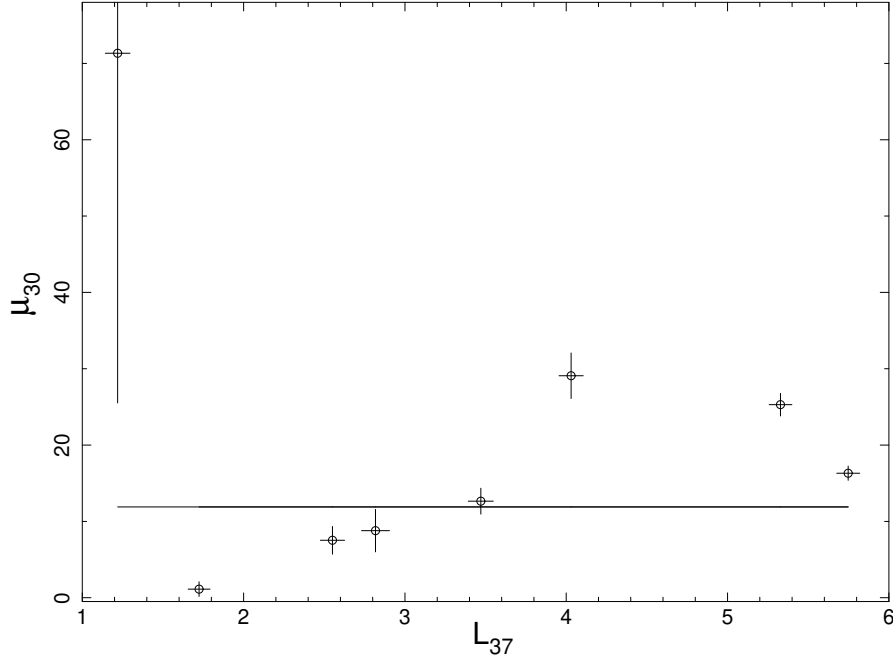


Figure 4.11: Plot of μ_{30} vs. L_{37} . The solid line is best fit with a constant which gives $\mu_{30} = 11.89$.

$$\dot{\nu}_{12} = 2.0n\zeta^{1/2}\mu_{30}^{2/7}R_6^{6/7}M_{1.4}^{-3/7}I_{45}^{-1}L_{37}^{6/7} \quad (4.1)$$

where $\dot{\nu}_{12}$, μ_{30} , R_6 , $M_{1.4}$, I_{45} are the spin frequency derivative, magnetic dipole moment, radius, mass and the moment of inertia of the neutron star given in the units of 10^{-12} Hz s $^{-1}$, 10^{30} G cm ^{3} , 10^6 cm, $1.4 M_{\odot}$, 10^{45} g cm ^{2} respectively. Also, the X-ray luminosity is termed as L_{37} in the unit of 10^{37} erg sec $^{-1}$. n and ζ are both dimensionless parameters. n is a function of “fastness parameter”(i.e. the ratio of the neutron star’s angular frequency to that of the disk at the radius r_0 at which the magnetic barrier terminates it). ζ is the factor by which r_0 differs from the Alfvén radius r_a (i.e. $r_0 = \zeta r_a$). According to Ghosh and Lamb (1979a;b) model, under slow-rotator condition $n \sim 1.39$ and $\zeta \sim 0.52$. Therefore, equation (4.1) reduces to (Sugizaki et al. 2017),

$$\dot{\nu}_{12} = kL_{37}^{\alpha} \quad (4.2)$$

where $k = 2.0\mu_{30}^{2/7}$ and $\alpha = 6/7 \simeq 0.857$.

For nominal values of $R_6 = M_{1.4} = I_{45} = 1$, measurements of the $\dot{\nu}_{12}$ vs. L_{37} provides

as a rough estimation of the magnetic dipole moment of the pulsar. Observationally, α ranges between 0.85-1. In Figure 4.10, the solid line is the best fit with $\alpha = 0.93$ and the dotted line is the fit for the theoretical value of α (i.e. 0.857). These two values are almost equal. We estimated the value of $k = 4.13 \pm 0.25$ and $\alpha = 0.93 \pm 0.04$ within 90% confidence range. From the best fit the $\dot{\nu}_{12}$ and L_{37} relation turns out to be

$$\dot{\nu}_{12} = 4.13 L_{37}^{0.93} \quad (4.3)$$

As, α is close to the theoretical value, we take

$$\Rightarrow 2.0 \mu_{30}^{2/7} = 4.13 \Rightarrow \mu_{30} = 12.6 \quad (4.4)$$

We also calculated the value of μ_{30} for each set of $\dot{\nu}_{12}$ & L_{37} and also the error on it from the propagated error of $\dot{\nu}_{12}$ & L_{37} . Figure 4.11 shows the variation in μ_{30} with luminosity. We modelled it with a constant for all the values of μ_{30} . The value of the constant came out to be 11.89 which is shown by the solid line in Figure 4.11.

4.5.3 Estimation of the surface magnetic field

The magnetic moment μ_{30} can be expressed in terms of surface magnetic field strength B_{12} and the radius of the pulsar R_6 as,

$$\mu_{30} = \frac{1}{2} B_{12} R_6^3 \phi(x) \quad (4.5)$$

where $\phi(x)$ is a correction factor (Wasserman and Shapiro 1983) on surface redshift parameter, $x (= \frac{Rc^2}{2GM})$ which is the ratio of the pulsar radius to the Schwarzschild radius. For a typical neutron star this ratio is ~ 2.4 and the correction factor $\phi(x) \sim 0.68$. So equation 4.5 becomes,

$$B_{12} = \frac{2 \times \mu_{30}}{0.68} \quad (4.6)$$

For $\mu_{30} \simeq 12.6$, $B_{12} \simeq 37$. So, the surface magnetic field strength estimated from the spin-up characteristics of GRO J2058+42 during its second outburst is $\simeq 37 \times 10^{12}$ Gauss. It is rather large compared to the surface magnetic field estimated from the cyclotron line in most accreting X-ray pulsars in Be X-ray binaries.

4.6 Discussion

In this chapter, we have studied the transient high mass Be X-ray binary pulsar GRO J2058+42. We have carried out timing and spectral analysis along with a study of its spin frequency evolution with the luminosity and estimated the strength of the magnetic field of the pulsar.

From the first *CGRO*-BATSE observation of the giant outburst (Wilson et al. 1998), pulse profiles of GRO J2058+42 were studied during the fast rise and then slower rise followed by the decay of the outburst and significant evolution in the pulse profile was witnessed from a fast rise to the decay. The pulse profile is representative of the beaming pattern. Hence it is a tool to observe any change in the accretion mode. The pulse profiles from the first outburst showed clockwise transformation of pulses. However, our phase aligned pulse profiles from the two *NuSTAR* observations, one during the rise and the other during the fall of the recent outburst look quite similar to each other (Figure 4.2 and 4.3). On the other hand, in terms of energy dependence, energy resolved pulse profiles from the previous and the recent outburst are in agreement. In the soft X-ray band there are at least four different pulse components at low energy which evolve to a single peaked profile at high energy (> 30 keV). Multiple peaks in low energy and single peaked pulses in higher X-ray energy band has also been seen in other pulsars like GRO J1008-57 (Naik et al. 2011b) and GX 304-1 (Devasia et al. 2011).

Previously, Wilson et al. (1998) studied the relationship between its spin-up rate and X-ray flux from its first outburst observation. The value of α (equation 4.2) was found to be ~ 1.2 , which is slightly higher than that is reported here for the second outburst. The X-ray luminosity shown in Figure 4.10 is calculated from the hard X-ray pulsed emission assuming that the X-ray spectral shape and pulsed fraction remained unchanged through out the period of Fermi observation. The medium energy X-ray light curve obtained with *MAXI*-GSC (2-20 keV), the hard X-ray light curve obtained with *Swift*-BAT (15-50 keV) and the pulsed flux of the source measured with *Fermi*-GBM during the outburst are shown in Figure 4.12. We have carried out separate analysis comparing the three light curves. We checked the hardness ratio using ratio of the count rates with time from *Swift*-BAT (15-50 keV) and *MAXI*-GSC (2-20 keV). We did not observe any significant change in the spectral shape. Also, we looked at the ratio of the count rate from the *Swift*-BAT and the pulsed flux from the *Fermi*-GBM observations during the entire

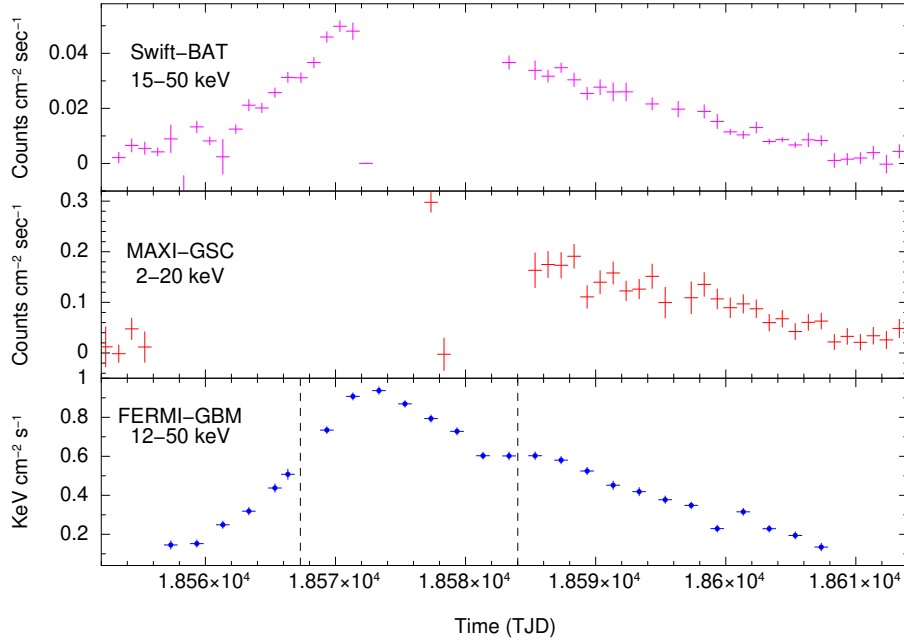


Figure 4.12: The lightcurve of GRO J2058+42 during it's second outburst observed by *Swift*-BAT in the top panel and by *MAXI*-GSC in the middle panel. The bottom panel shows the pulsed fluxes observed by *Fermi*-GBM and the two black vertical lines represent the start time of the two *NuSTAR* observation.

outburst. The ratio was almost equal during the entire period, which indicated that there was no significant change in the pulsed fraction during this period. Within error bars the X-ray spectral shape and hard X-ray pulse fraction remained the same during the outburst. Another unknown factor that may contribute to the pulse period and period derivative measurements is the orbital motion of the pulsar. As the orbital parameters of GRO J2058+42 are unknown, it is not possible to correct the data for orbital motion.

From the $\dot{v}_{12}$ vs. L_{37} relationship, we calculated the value of the magnetic moment of the neutron star and from that we estimated the magnetic field strength which came out to be $\simeq 37 \times 10^{12}$ Gauss. The high magnetic field in neutron stars causes the formation of the CRSF in the X-ray spectra. This broad absorption like feature in the energy spectrum, arises due to the X-ray photons in the accretion column scattered by the plasma electron quantized in the Landau levels. The magnetic field strength and the fundamental cyclotron energy is related as $B_{12} = \frac{1}{\sqrt{1-x^2}} \frac{E_{cyc}}{11.6 \text{ keV}}$. For Be XRPs the CRSF is observed within a energy range of 12–76 keV (Sugizaki et al. 2017). The energy spectra of GRO J2058+42 does not show any CRSF in 3-78 keV energy band. From the estimated $B_{12} \simeq 37$,

we get $E_{cyc} \simeq 328$ keV. So, the cyclotron line is expected to be outside the energy band of the X-ray spectrum measured with *NuSTAR*.

The B_{12} and E_{cyc} estimated for GRO J2058+42 from the accretion torque is much larger than that of other Be XRBs. One of the reason can be an under estimation of the total X-ray luminosity. While estimating the total flux from the pulsed flux, we considered the energy range of *NuSTAR* i.e. 3-78 keV. So, the bolometric X-ray luminosity in 0.1-100 keV band should be higher. With the best fit *NuSTAR* model, for both the observations, we calculated the 0.1-100 keV flux, which is larger than the 3-78 keV flux by only 3-4%. However, the surface magnetic field is high enough that we do not see the CRSF in the *NuSTAR* band, which indicates to a $B_{12} \geq 9$ and it is higher than the magnetic fields confirmed for HMXBs (Caballero and Wilms 2012). There are few other Be XRBs like EXO 2030+375 (Naik et al. 2013), 2S 1417-624 (Gupta et al. 2018) and GS 0834-430 (Miyasaka et al. 2013), for which no CRSF feature has been found. So, it can be a possibility that some of the Be XRBs may have a high magnetic field of the order of 10^{13} Gauss. Despite a non-detection of any cyclotron line in the spectra, we are not entirely rejecting the idea of presence of it and provided an upper limit of the optical depth for each energy (Figure 4.9). In this context, we would like to mention Molkov et al. (2019). They reported a cyclotron line at 10 keV along with its two harmonics at a particular pulse phase. Our study shows that we do not find any cyclotron line in the phase averaged spectra which is in agreement with their study and along with non-detection of any feature, it is customary to report the upper limit of the same.

Along with it, we also looked at the difference in the estimation of the strength of the surface magnetic field from spin-up rate and luminosity relation, and from the observed cyclotron line. We calculated the μ_{30} and B_{12} from the proportionality factor k in the spin-up rate and luminosity relation in equation 4.1 provided in Sugizaki et al. (2017) for nine transient pulsars for which cyclotron line has been observed. Canonical values of $R_6 = 1$, $M_{1.4} = 1$, $I_{45} = 1$ for neutron stars have been assumed for the calculation. We also calculated the magnetic field value from the observed cyclotron line energy using the relation $B_{12} = \frac{1}{\sqrt{1-x^{-1}}} \frac{E_{cyc}}{11.6 \text{ keV}}$ and compared them. Table 4.4 represents a large difference in these two B_{12} values. The two estimates of magnetic field strength from torque-luminosity relation and cyclotron line energy differ by large factors, from 10^{-5} to 4.0. Part of these differences may arise if R_6 , $M_{1.4}$, I_{45} , beaming factor, and effect of misalignment between spin axis and magnetic axis etc. for each individual source is different from 1.0. Another

Table 4.4: Comparison of estimated strength of surface magnetic field of nine transient pulsars from the k^* and from observed cyclotron energy

Source	B_{12} from k^*	B_{12} from observed E_{cyc}
4U 0115+63	1.15×10^{-2}	1.81
X 0331+53	3.76×10^{-5}	3.50
RX J0520.5-6932	6.07×10^{-2}	3.56
H 1553-542	5.21×10^{-1}	3.08
XTE J1946+274	1.23	3.95
KS 1947+300	5.33	1.38
GRO J1008-57	1.41×10^{-1}	8.58
A 0535+262	8.03×10^{-1}	5.28
GX 304-1	2.51×10^{-1}	6.06

factor is the uncertainty in source distance and hence on luminosity. We note here the strong dependence of the magnetic moment on radius of the neutron star (third order) and distance (sixth order). A difference in distance by a factor of 1.5 can result in a factor of ~ 10 difference in estimation of the magnetic field strength. Therefore, the actual surface magnetic field strength of GRO 2058+42 can be different from its estimated value from spin-up rate and luminosity relation.

After the second giant outburst, the X-ray luminosity of GRO J2058+42 dropped significantly leading to the non-detection of the pulsar with *Fermi*-GBM. It is a common phenomena for Be X-ray binaries and indicative of the onset of the propeller regime when the magnetic field inhibits the accretion process (Illarionov and Sunyaev 1975, Lutovinov et al. 2017, Revnivtsev and Mereghetti 2015, Tsygankov et al. 2016). After the outburst, GRO J2058+42 was observed again by *NuSTAR*. The visibility of BeXRBs during the quiescent state can be explained by various emission mechanisms like : (a) accretion from

a cold disk, (b) penetration of some of the accretion matter into the magnetosphere in the propeller regime, (c) cooling of the neutron star which was heated up by the accreted matter during the outburst (i.e. the deep crustal heating model), etc (Tsygankov et al. 2017b and the references there in). The *NuSTAR* observation (MJD 58723.53) showed clear pulsations in the source which suggests the presence of a compact emission region. The strong pulsations, and the pulse profiles in different intensity levels suggest against the low luminosity state to be from the heated neutron star surface. The spectra (FRMA and FPMB) are well described with a power-law which suggests a non-thermal emission. The spectral parameters are different in the low luminosity state than during the outburst. The photon index has increased almost by a factor of two causing a softer shape of the energy spectra. The detected flux is almost 10% of what was observed from the *NuSTAR* observations during the outburst. The estimated luminosity from the observed flux is $\sim 3 \times 10^{36} \text{ erg s}^{-1}$.

The limiting accretion luminosity on the onset of the propeller regime is given by (Campana et al. 2002),

$$L_{lim} = 4 \times 10^{37} \zeta^{7/2} B_{12}^2 P_s^{-7/3} M_{1.4}^{-2/3} R_6^5 \text{ erg s}^{-1} \quad (4.7)$$

To calculate L_{lim} for GRO J2058+42, we used the estimated value of $B_{12} = 37$ (magnetic field strength in the unit of 10^{12} Gauss) and $P_s = 193.37$ (spin period in seconds). From equation 4.7, the limiting accretion luminosity is $\sim 1.5 \times 10^{32} \text{ erg s}^{-1}$. It is much smaller than the observed luminosity of the source. L_{lim} will be even smaller for a lower value of the magnetic field strength (order of 10^{12} Gauss) of the neutron star. It is commonly believed that a Be X-ray pulsar enters the propeller regime once its luminosity falls below L_{lim} . The luminosity of GRO J2058+42 indicates that the pulsar has not entered into a propeller regime but is possibly accreting from a cold recombined accretion disc, which is the usual case for pulsars with longer spin periods (dozens of seconds) (Tsygankov et al. 2017a).

Chapter 5

A study of the pulse shape of persistent accreting pulsars at different accretion torque



5.1 Introduction

Spin evolution of accreting pulsars is a consequence of the transfer of angular momentum via accretion and the disk-magnetosphere interaction of the neutron star. According to the standard model (Ghosh and Lamb 1979b), the accretion torque is expected to be related to the mass accretion rate and hence to the bolometric X-ray luminosity (Bildsten et al. 1997). This correlation is quite evident in the case of Be X-ray binaries like EXO 2030+375, GRO J2058+42 etc. (Wilson et al. 2002, Sugizaki et al. 2015, Sugizaki et al. 2017, Kabiraj and Paul 2020). Steady long-term spin up and spin down episodes are seen in some of the persistent sources like 4U 1626-67 and GX 1+4 (Chakrabarty et al. 1997, Beri et al. 2014). A possible explanation of such torque behavior can be the transition between prograde and retrograde disk. In the presence of a stable prograde accretion disk, the accreted matter will provide a positive torque and the neutron star will spin up. A retrograde disk will feed negative torque to the neutron star causing it to spin down. In general, the persistent disk-fed systems go through steady spin evolution with longer time scales (for years) and the transfer of the angular momentum is interpreted based on the disk behaviour. A correlation or an anti-correlation between the spin up rate and luminosity is expected depending on a prograde or a retrograde disk respectively.

However, in the case of wind fed persistent sources the process of the angular momentum transfer via accretion is complex. Some of these sources show erratic torque reversals without any clear correlation with the X-ray luminosity (5.1) (Nelson et al. 1997, Torkelsson 2003). The long term X-ray flux history and the spin frequency history of these sources indicate that mass accretion cannot be the only mechanism behind the transfer of the torque. In one of the pulsar 4U 1626-67, the shape of the pulse profile was found to be strongly correlated with the sign of the accretion torque (5.2) (Beri et al. 2014). The pulse profiles reflect the intensity variation as the neutron star rotates. The shape of the profile depends on multiple factors like angle between the spin and magnetic axes, the accretion and emission region, gravitational bending of light, and the resultant emission beam which is energy dependent.. The observed correlation in 4U 1626-67 suggests that the pulse shapes at different accretion torque bear the information of the change in the accretion mode. In this work, we have used the pulse profiles of four wind-fed pulsars as a tool to observe any changes in the accretion geometry at different accretion torque.

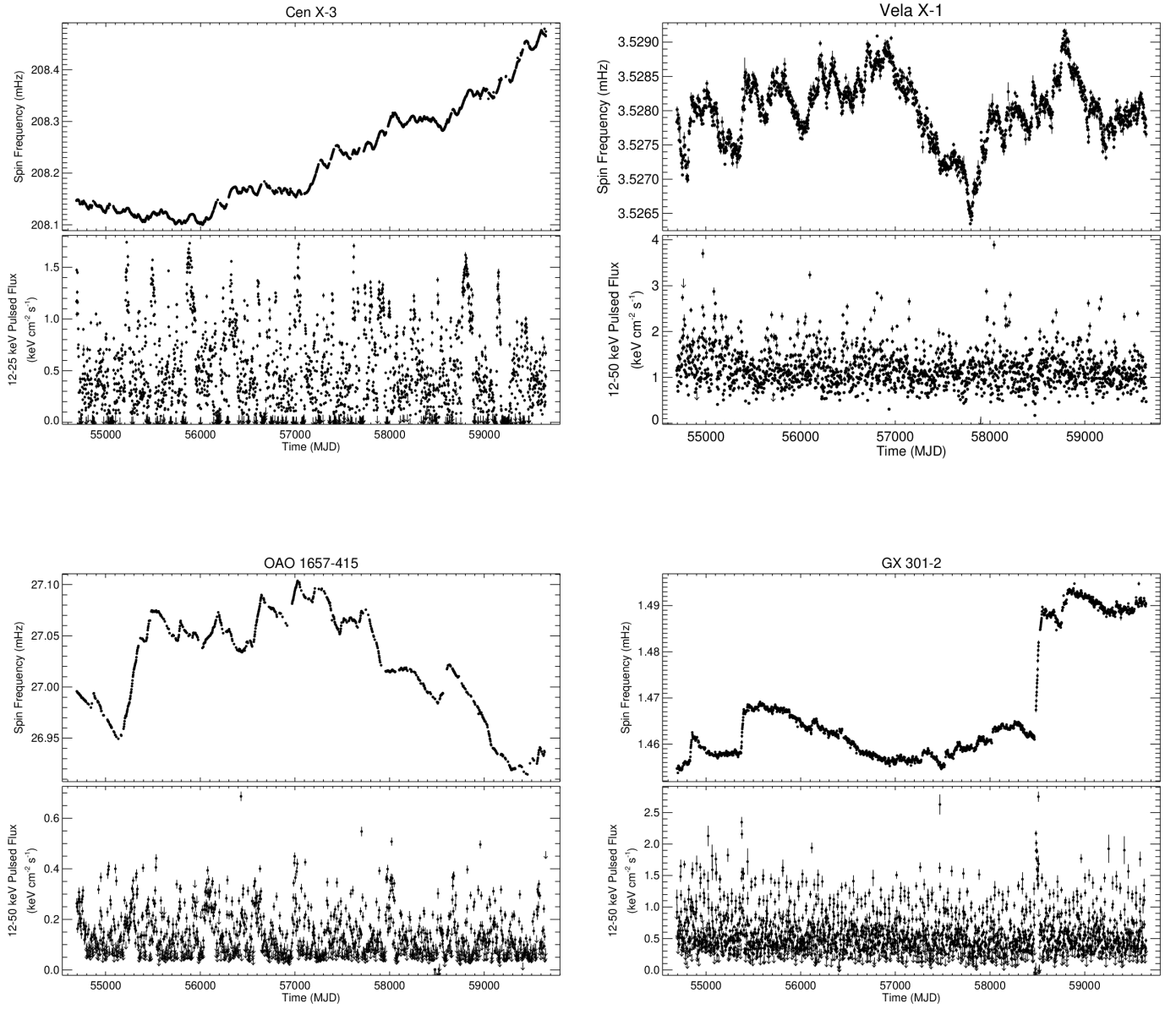


Figure 5.1: The long-term spin evolution of four persistent pulsars that show erratic spin-up and down behaviour.

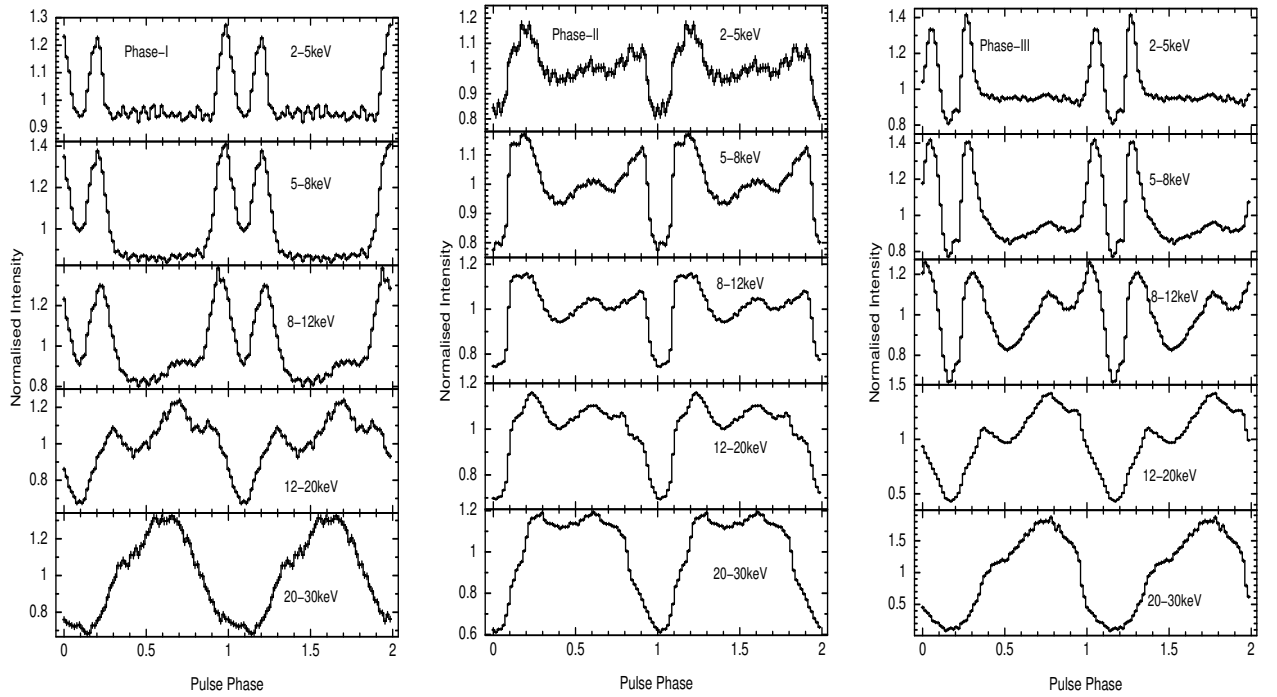


Figure 5.2: Energy resolved pulse profile of 4U 1626-67 in three phases of accretion torque. The profiles in the first panel on the left is during the spin-up era between 1977 and 1990. The middle panel represents the profiles during the spin-down era between 1990 and 2008. The panel on the extreme right represents the profiles during the spin-up era after 2008. From the top panel, the energy ranges are 2-5 keV, 5-8 keV, 8-12 keV, 12-20 keV and 20-30 keV. This plot is taken from [Berl et al. \(2014\)](#).

5.1.1 Cen X-3

Cen X-3 was discovered in 1967 by a rocket-based detector (Chodil et al. 1967). Later *Uhuru* observation revealed its binary and pulsating nature (Giacconi et al. 1971a, Schreier et al. 1972), and it became the first identified accretion powered pulsar. The orbital period of Cen X-3 is 2.1 days and the orbit has a low eccentricity of 0.0016 with the projected semi-major axis $a_x \sin i = 39.56 \pm 0.07$ lt-sec (Nagase et al. 1992b, Bildsten et al. 1997). It is an eclipsing persistent X-ray pulsar with a spin period of ~ 4.8 seconds and mass of $1.21 \pm 0.21 M_\odot$ (Schreier et al. 1972, Ash et al. 1999). The companion is an O-type supergiant star V779 Cen (Krzeminski 1974) with a radius and mass of $\sim 12 R_\odot$ and $\sim 20.5 \pm 0.7 M_\odot$ (Hutchings et al. 1979, Rappaport et al. 1983, Ash et al. 1999). The distance to the source was estimated to be 8 kpc (Krzeminski 1974), although Thompson and Rothschild (2009) gave a value of 5.7 ± 1.5 kpc. The X-ray spectrum of Cen X-3 is described by an absorbed power law with multiple spectral features like : three iron emission lines at 6.4 keV, 6.67 keV and 6.97 keV, a ~ 10 keV bump and a Cyclotron Scattering Resonance Feature at ~ 28 keV (Santangelo et al. 1998, Coburn et al. 2002, Suchy et al. 2008, Naik and Paul 2012). It is a bright pulsar with luminosity of $\sim 5 \times 10^{37} \text{ ergs}^{-1}$ indicating accretion via an accretion disk (Suchy et al. 2008). The optical light curve also indicates accretion via Roche Lobe overflow (Tjemkes et al. 1986). Even if disk-fed, Cen X-3 does not show any correlation between its spin frequency derivative and luminosity (Tsunemi et al. 1996, Raichur and Paul 2008a,b). The average long-term luminosity varies by a factor of ~ 40 . It has been suggested that the wind-fed mass transfer is also present (Walter et al. 2015, El Mellah et al. 2019).

5.1.2 Vela X-1

Vela X-1 (a.k.a. 4U 0900-40) is an eclipsing wind-fed HMXB with a neutron star situated at a distance of $2.42^{+0.19}_{-0.17}$ kpc (Bailer-Jones et al. 2018). It was detected by Chodil et al. (1967) and has a moderate X-ray luminosity of $\sim 4 \times 10^{36} \text{ ergs}^{-1}$ (Nagase et al. 1986). The binary orbit has a radius $\sim 53.4 R_\odot$ (van Kerkwijk et al. 1995a) and low eccentricity of $e \approx 0.09$ with an inclination of $i > 73^\circ$ with $a_x \sin i = 112.98 \pm 0.35$ lt-sec (Joss and Rappaport 1984, Deeter et al. 1987, Boynton et al. 1986). The neutron star has a pulse period of 283 seconds and mass between $1.8 M_\odot$ to $2.3 M_\odot$ (Barziv et al. 2001, Quaintrell et al. 2003, Rawls et al. 2011). The optical companion is HD 77581, which is a B0.5Ib-type supergiant with a radius of $30 R_\odot$ and a mass of $23.5 M_\odot$ (Brucato and Kristian 1972, Hiltner

et al. 1972, Forman et al. 1973, van Kerkwijk et al. 1995b, Quaintrell et al. 2003, Rawls et al. 2011). The orbital period is ~ 8.94 days (Barziv et al. 2001). The X-ray spectrum is described by a power law with high energy cut off and a soft excess below 3 keV (Nagase et al. 1986, Haberl and White 1990, Kreykenbohm et al. 1999). The iron $K\alpha$ emission line at ~ 6.42 keV and iron absorption edge at ~ 7.24 keV are present in the spectrum (Nagase et al. 1986, Watanabe et al. 2006). The cyclotron resonance scattering features were reported around 25-32 keV and at around ~ 55 keV (Orlandini et al. 1998, Kendziorra et al. 1992, Choi et al. 1996, Kreykenbohm et al. 2002, La Barbera et al. 2003, Attié et al. 2004). Kreykenbohm et al. (2002) confirms that the fundamental cyclotron line exists at ~ 25 keV and the line at ~ 50 keV is the first harmonic. Vela X-1 has an average X-ray luminosity of $\sim 4 \times 10^{36} \text{ erg s}^{-1}$ which varies by a factor of 20-30 (Kreykenbohm et al. 2008). The spin frequency evolution of Vela X-1 does not have a steady long-term trend, but it shows sporadic spin-up and spin-down episodes. This behaviour is attributed to the presence of a transient disk around the neutron star (El Mellah et al. 2019, Fryxell and Taam 1988).

5.1.3 OAO 1657-415

The HMXB OAO 1657-415 was discovered as an X-ray source by the Copernicus satellite (Polidan et al. 1978) and was recognized as a ~ 38.22 second pulsar by White and Pravdo (1979). The CGRO *BATSE* observation determined the orbital period of ~ 10.4 days with an eclipse of ~ 1.7 days (Chakrabarty et al. 1993). The binary orbit has an eccentricity of $e \sim 0.104 = 0.005$ with the projected semi-major axis $a_x \sin i \sim 104.0 \pm 0.5$ lt-sec (Chakrabarty et al. 1993). From these orbital parameters and assuming the mass of the neutron star $= 1.4M_{\odot}$, the radius and mass of the companion was calculated in the range of $25 - 32R_{\odot}$ and $14 - 18M_{\odot}$ (Chakrabarty et al. 1993). Mason et al. (2009) suggest that the companion is an Ofpe/WN9-type supergiant, which is in a transition phase between an OB and Wolf-Rayet star. The distance of the binary was estimated to be 7.1 ± 1.3 kpc by Audley et al. (2006) from the ASCA observation of its dust scattered X-ray halo. Assuming a distance of 2.2 kpc (Malacaria et al. 2020), Saavedra et al. (2022) estimated the average unabsorbed X-ray luminosity in 3-79 keV energy band to be $\sim 10^{35} - 10^{36} \text{ erg s}^{-1}$. The X-ray spectrum is highly absorbed with a hydrogen column density of $\sim 10^{23} \text{ atom cm}^{-2}$ and can be described by a cutoff power law model (Parmar et al. 1980, Orlandini et al. 1999, Pradhan et al. 2014). It has iron emission lines at 6.4 keV, 6.7, 6.7 and 7.1 keV and a possible cyclotron absorption feature at ~ 36 keV (Orlandini

et al. 1999, Chakrabarty et al. 2002, Audley et al. 2006, Pradhan et al. 2014, Pradhan et al. 2019). Pradhan et al. (2019) reported that a Compton shoulder around ~ 6.3 keV is present due to the scattering of the iron fluorescence photons. OAO 1657-415 went through several spin up and spin down episodes (White and Pravdo 1979, Chakrabarty et al. 2002) and the spin frequency evolution does not correlate with the X-ray luminosity (Baykal 2000). To describe the spin-up and spin-down trend of OAO 1657-415, Jenke et al. (2012) explained that the accretion in OAO 1657-415 happens both via a transient disk and stellar wind.

5.1.4 GX 301-2

GX 301-2 (a.k.a. 4U 1223-62) was discovered in 1975 by *Ariel5* by White et al. (1976). The binary has a 41.5 day eccentric orbit ($e \approx 0.47$) with $a_x \sin i = 368.3 \pm 3.7$ lt-sec (Sato et al. 1986, Koh et al. 1997). This binary system harbours a neutron star with a pulse period of ~ 676 seconds (Nagase 1989). The supergiant companion Wray-977 is a B1 Ia+ star with a mass between 39 - 53 $M_\odot$ with a radius of 62 $R_\odot$ (Vidal 1973, Parkes et al. 1980, Kaper et al. 1995, Kaper et al. 2006). It is situated at a distance of 3.1 kpc (Hammerschlag-Hensberge et al. 1979, Kaper et al. 2006). GX 301-2 exhibits X-ray flaring activity about 1.4 days before the periastron passage. During these pre-periastron flares the X-ray luminosity increases by a factor of ~ 25 (Pravdo et al. 1995, Sato et al. 1986). The overall X-ray luminosity varies within the range of $(2-200) \times 10^{35} \text{ erg s}^{-1}$ (Endo et al. 2002). The X-ray spectrum is heavily absorbed and the column density varies with the orbital phase. This variation is possibly caused by the clumps in the stellar wind (Pravdo and Ghosh 2001, Mukherjee and Paul 2004, Leahy and Kostka 2008). The spectrum is well described by a power law with a high energy cutoff (Watanabe et al. 2003, Mukherjee and Paul 2004, La Barbera et al. 2005, Suchy et al. 2012). It has a prominent Fe $K\alpha$ emission line at 6.4 keV and cyclotron resonant scattering feature around ~ 37 keV (Makino et al. 1985, Mihara 1995, Watanabe et al. 2003). GX 301-2 also shows a Compton shoulder of the Fe- $K\alpha$ line extending to ~ 6.24 keV (Watanabe et al. 2003). The pulsar alternates between spin-up and spin-down with random walk variation (de Kool and Anzer 1993). However, it also goes through prolonged rapid spin-up episodes. Three of such rapid spin-up episodes were monitored by *Fermi*-GBM since its launch during MJD 54825-54860, MJD 55367-55403 and MJD 58465-58567. These episodes lasting for over tens of days takes place probably due to the accretion from a transient disk.

All of these four pulsars show erratic spin up and down behaviour. We have examined

the pulse profiles from *Fermi*-GBM and investigated any possible correlation between the pulse shapes and the local (w.r.t. time) spin frequency change rate. In the next section, we have discussed about the data and observation. Lastly, we have presented the implications of our result.

5.2 Observation

5.2.1 Data

For this work, we received energy resolved pulse profiles from *Fermi*-GBM team for all the four pulsars at every ~ 100 days gap since the launch of *Fermi* mission. The time for each observations are tabulated. Each data set contained pulse profiles in three different energy bands (i.e. 8-12 keV, 12-25 keV and 25-50 keV). These composed of ‘pulse phase’, ‘amplitude’ and ‘error on amplitude’. Some of the data points had no associated errors with amplitude. They represent the superposition of the cosine and sine functions used in the profile.

We have also used the pulse frequency, pulsed flux and its errors as determined by the Fermi GBM team. The method of measuring the pulse frequency and the determination of the error is reported in [Malacaria et al. \(2020\)](#). In the following paragraph, we have provided a brief description of the method used and we refer to this paper for more details.

For a particular observed event from a pulsar, there is a epoch time of the observation. After the detection of an event at a particular epoch time, at first the barycenter correction is done. The corrected epoch time is achieved taking into account various factors like Einstein delay, Shapiro delay, Roemer delay and the clock correction for differences between the observatory clocks and terrestrial time standards. Now, because of the orbital motion of a pulsar, the arrival time of a pulse is Doppler shifted. When the orbital parameters are known, the corrected epoch time is then further corrected for orbital motion of the neutron star to get the actual pulse emission time. Fermi continuously monitors the pulsars and tracks the arrival time of individual pulses. A periodic pulsar signal can be described with Fourier harmonic series. The background subtracted source count rates are fit with a constant plus a Fourier expansion to determine a pulse profile. The profiles are divided into six day intervals and the pulse frequency and mean profile determined in each interval with a search of pulse frequency for the maximum of the Y_n ($n=2$) statistic

(Camero-Arranz et al. 2009, Malacaria et al. 2020) (Camero-Arranz 2010; Finger et al. 1999, Malacaria et al. 2020). Y_n is the F statistic summed for each Fourier component in which the search is performed and used as a tool to find pulse frequency from a series of pulse profiles. All the errors associated with the pulse frequency or the pulsed flux are calculated from the Hessian of the Y_n matrix. In order to achieve epoch times and pulse frequencies along with the associated uncertainties, this entire analysis and correction is done by Fermi team and then we have used these products for our science goal.

5.2.2 The measurement of spin change rate and the pulsed flux

The spin change rate at a particular time was estimated following the similar procedure described in 4.5.1. We have taken 10 days on each side of a particular time window and fitted the spin frequencies ν_s in those intervals w.r.t. time with a linear function. Our objective was to select a small time interval, but also have enough number of ν_s measurements to look for the local change in the spin frequency when a pulse profile was extracted. We took a minimum of 5 ν_s measurements to obtain a $\dot{\nu}_s$. For some of the observations, we had less than 5 ν_s values when we took 10 days on each side of the time windows. For those cases, we have expanded the interval by tens of days to achieve minimum number of ν_s s to estimate the local $\dot{\nu}_s$. In the Tables 5.1, 5.2, 5.3, 5.4, 5.5, 5.6 and 5.7, we have provided the start and end time of the time windows when the pulse profiles were extracted and the time span that has been used to determine the $\dot{\nu}_s$ s. For the pulsed flux, we have simply taken its observed value between T_{start} and T_{end} from the *Fermi*-GBM long-term observation. Figure 5.3 shows the plots of the spin change rate with pulsed flux measurements for the four sources.

5.2.3 Effect of decoherence timescales

The pulse profiles are periodic signals over time. In order to get a proper signal, integration of weak signals over a significant duration of time is required. The sampling of the weak signals should be done in such manner so that it preserves the periodic pulse phase coherence. Signal decoherence occurs while doing pulse-timing observations if the pulses do not arrive at equal time intervals. This happens for mainly two reasons - 1) the periodic Doppler shift due to binary orbital motion, 2) the intrinsic change in the spin frequency of the pulsar. This decoherence leads to the nondetection of the weak signals and so reduces the sensitivity which results in smearing of the pulse profiles. The pulse

arrival times of the *Fermi* observations are corrected for the orbital Doppler shift before the extractions of the pulse profiles. Therefore, we have to consider the signal decoherence only due to the intrinsic spin frequency change of the pulsars. In this work, we are comparing the shape of the pulse profiles at different accretion torque. If the intrinsic spin change rate is higher than the sampling time over which the pulses are accumulated to construct the profile, then the profile will lose its original shape. For this, let us consider that the pulsar has a steady spin frequency change rate of $\dot{\nu}$ during the time of the observation. If the pulse phase at the start of the observation was ϕ_o , then after a time t the pulse phase ϕ should be,

$$\phi = \phi_o + \nu t + \frac{1}{2} \dot{\nu} t^2 + \dots \dots \dots \text{(ignoring the higher order terms)} \quad (5.1)$$

According to the definition the decoherence is when the phase shift $\Delta\phi = \frac{1}{2}$. Therefore, the decoherence timescale τ is given by the condition,

$$\Delta\phi = \frac{1}{2} = \frac{1}{2} \dot{\nu} \tau^2 \Rightarrow \tau = \sqrt{\frac{1}{\dot{\nu}}} \quad (5.2)$$

So, we calculated the decoherence timescales from the spin change rates and compared it with the time of the observations. For the calculation of the decoherence timescale for an observation, we took the modulus values of the $\dot{\nu}$ for each observation. We have also provided the value of $\frac{1}{2} \dot{\nu} T^2$ (T = time of observation to extract a pulse profile) for each observation to know the amount of the decoherence associated with the pulse profiles. All these values from these calculations are tabulated in Table 5.1, 5.2, 5.3, 5.4, 5.5, 5.6 and 5.7.

We notice that for some observations of OAO 1657-415 the contribution of the decoherence ($\frac{1}{2} \dot{\nu} T^2$) is almost 10% – 20%. Taking this fact into account, it is possible that the pulse profiles might have lost their true shape because of the distortion. Therefore, we did not analyze this source further and continued our study with the other three pulsars.

5.2.4 Pulse profiles

Figure 5.4, 5.5, 5.6 are the 12-25 keV pulse profiles of Cen X-3, Vela X-1 and GX 301-2 with corresponding spin periods. These profiles are from the observations that are evenly distributed over time since the launch of the *Fermi* mission. The profiles are plotted by normalizing the average of the pulse profile to ‘0’. We sorted the profiles in terms of

the decreasing order of the spin frequency derivative. The different colours are for the different times of the observations.

Also, we took the 12-25 keV pulse profiles of GX 301-2 during the three rapid spin-up episodes during MJD 54825-54860 (first), MJD 55367-55403 (second) and MJD 58465-58567 (third) and repeated the same procedure. For the third outburst, we also had data for another 30 days after the outburst, we looked into that too.

In the figure 5.4, we see that pulse profiles of Cen X-3 are single peaked and we have aligned the profiles with respect to the peak. They have similar shapes except for the variation in the amplitude. The four panels show pulse profiles during four different spin change rates. In this Figure, in the top left the range of $\dot{\nu}$ is from 6.50×10^{-7} to $2.96 \times 10^{-7} \text{ Hz/day}$, in the top right the range of $\dot{\nu}$ is from 2.84×10^{-7} to $5.59 \times 10^{-8} \text{ Hz/day}$, in the bottom left the range of $\dot{\nu}$ is from 4.08×10^{-8} to $-2.88 \times 10^{-7} \text{ Hz/day}$, and lastly, in the bottom right the range of $\dot{\nu}$ is from -2.96×10^{-7} to $-4.80 \times 10^{-7} \text{ Hz/day}$.

Similarly, in the Figure 5.5, the pulse profiles of Vela X-1 are shown. All the profiles have two almost equally strong peaks and are aligned with respect to its main minima. The overall shape of the profiles do not change with the spin frequency change rate. In this Figure, in the top left the range of $\dot{\nu}$ is from 2.65×10^{-8} to $7.07 \times 10^{-9} \text{ Hz/day}$, in the top right the range of $\dot{\nu}$ is from 6.70×10^{-9} to $6.52 \times 10^{-10} \text{ Hz/day}$, in the bottom left the range of $\dot{\nu}$ is from 4.58×10^{-10} to $-5.35 \times 10^{-9} \text{ Hz/day}$, and lastly, in the bottom right the range of $\dot{\nu}$ is from -5.62×10^{-9} to $-1.63 \times 10^{-8} \text{ Hz/day}$.

Figure 5.6 shows the pulse profiles of GX 301-2. These are the profiles from the observations that are uniformly distributed over time since 2008 with about ~ 100 days gap between each observation. GX 301-2 mostly goes through irregular spin-up and down, and these profiles are the representative of the beaming pattern during these times. They are double peaked and one of them is weaker. We have aligned the profiles with respect to the higher peak. The amplitude of the profiles vary from one to another. We have also presented the pulse profiles during the three rapid spin-up episodes (Figure 5.7 and 5.8) and arranged them in the decreasing order of the spin frequency change rate. They are also double peaked and the relative strength of these two peak vary. However, all the pulse profiles of GX 301-2 have similar shapes.

Table 5.1: Time of observations, spin change rate and decoherence time scale of Cen X-3

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
54690.02	54691.02	1.00	10	1.38e-07	9.15e+00	5.97e-03
54795.85	54796.99	1.14	10	-3.82e-07	5.51e+00	2.13e-02
54891.86	54893.47	1.61	10	-2.96e-07	6.25e+00	3.31e-02
54992.02	54993.65	1.63	10	3.71e-10	1.77e+02	4.25e-05
55090.11	55091.74	1.63	10	-3.54e-07	5.72e+00	4.06e-02
55190.29	55191.91	1.62	10	-2.10e-07	7.43e+00	2.38e-02
55290.46	55292.08	1.62	10	-9.96e-08	1.08e+01	1.12e-02
55390.65	55392.25	1.60	10	6.51e-08	1.33e+01	7.21e-03
55490.82	55492.43	1.61	10	2.96e-07	6.25e+00	3.31e-02
55590.99	55592.59	1.60	10	-3.18e-07	6.03e+00	3.53e-02
55693.25	55694.88	1.63	10	-1.10e-07	1.02e+01	1.27e-02
55791.34	55792.96	1.62	10	-3.44e-07	5.80e+00	3.91e-02
55891.52	55893.14	1.62	10	1.26e-07	9.58e+00	1.43e-02
55991.69	55993.32	1.63	10	-1.99e-07	7.62e+00	2.29e-02
56091.86	56093.46	1.60	10	1.39e-07	9.12e+00	1.54e-02
56206.66	56208.28	1.62	10	-2.78e-07	6.45e+00	3.16e-02
56298.49	56300.10	1.61	10	6.50e-07	4.22e+00	7.31e-02
56390.33	56391.93	1.60	10	-4.49e-07	5.08e+00	4.97e-02
56490.48	56492.11	1.63	10	2.15e-07	7.33e+00	2.47e-02
56592.79	56594.36	1.57	10	7.63e-08	1.23e+01	8.13e-03
56690.83	56692.46	1.63	10	-4.80e-07	4.91e+00	5.53e-02
56791.01	56792.62	1.61	10	2.84e-07	6.38e+00	3.18e-02
56891.18	56892.81	1.63	10	5.59e-08	1.44e+01	6.41e-03
56991.37	56992.98	1.61	10	-2.23e-07	7.21e+00	2.49e-02
57097.83	57099.41	1.58	10	4.08e-08	1.68e+01	4.41e-03
57191.75	57193.33	1.58	10	5.35e-07	4.65e+00	5.76e-02
57291.90	57293.46	1.56	10	-3.81e-07	5.51e+00	4.02e-02
57392.06	57393.68	1.62	10	6.16e-07	4.33e+00	6.99e-02
57490.14	57491.76	1.62	10	-2.88e-07	6.34e+00	3.27e-02
57602.84	57604.47	1.63	10	6.47e-07	4.23e+00	7.41e-02
57723.88	57725.51	1.63	10	5.48e-07	4.60e+00	6.31e-02
57790.67	57792.29	1.62	10	-1.00e-07	1.07e+01	1.13e-02

Table 5.1: Time of observations, spin change rate and decoherence time scale of Cen X-3 (continued)

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
57890.84	57892.46	1.62	10	4.40e-07	5.13e+00	5.00e-02
57991.01	57992.61	1.60	10	5.85e-07	4.45e+00	6.45e-02
58091.19	58092.82	1.63	10	-3.44e-07	5.80e+00	3.94e-02
58191.36	58192.97	1.61	10	2.23e-07	7.21e+00	2.49e-02
58291.53	58293.14	1.61	10	1.48e-07	8.85e+00	1.65e-02
58393.80	58395.43	1.63	10	-6.31e-08	1.35e+01	7.23e-03
58500.25	58501.86	1.61	10	-3.30e-07	5.92e+00	3.69e-02
58592.05	58593.68	1.63	10	5.25e-07	4.69e+00	6.02e-02
58690.15	58691.77	1.62	10	-3.55e-07	5.71e+00	4.03e-02
58790.34	58791.95	1.61	10	4.26e-07	5.21e+00	4.77e-02

Table 5.2: Time of observations, spin change rate and decoherence time scale of Vela X-1

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
54690.01	54692.04	2.03e+00	10	-3.22e-09	5.99e+01	5.72e-04
54790.68	54793.03	2.35e+00	10	-1.63e-08	2.66e+01	3.90e-03
54893.51	54895.87	2.36e+00	10	-2.32e-09	7.07e+01	5.57e-04
54992.11	54994.47	2.36e+00	10	-1.01e-09	1.07e+02	2.44e-04
55090.74	55092.98	2.24e+00	10	9.50e-10	1.10e+02	2.06e-04
55191.73	55194.08	2.35e+00	10	-1.63e-08	2.66e+01	3.89e-03
55290.31	55292.67	2.36e+00	10	6.07e-09	4.37e+01	1.46e-03
55391.30	55393.61	2.31e+00	10	7.24e-09	4.00e+01	1.67e-03
55494.12	55496.47	2.35e+00	10	7.98e-09	3.81e+01	1.90e-03
55592.72	55595.09	2.37e+00	10	4.86e-09	4.88e+01	1.18e-03
55691.33	55693.69	2.36e+00	10	2.31e-08	2.24e+01	5.58e-03
55792.34	55794.68	2.34e+00	10	-3.03e-09	6.18e+01	7.18e-04

Table 5.2: Time of observations, spin change rate and decoherence time scale of Vela X-1 (continued)

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
55890.95	55893.29	2.34e+00	10	-5.62e-09	4.54e+01	1.33e-03
55991.94	55994.28	2.34e+00	10	-5.75e-09	4.49e+01	1.36e-03
56090.52	56092.86	2.34e+00	10	1.09e-08	3.25e+01	2.58e-03
56193.34	56195.72	2.38e+00	10	2.65e-08	2.09e+01	6.48e-03
56291.94	56294.28	2.34e+00	10	7.09e-09	4.04e+01	1.68e-03
56390.55	56392.92	2.37e+00	10	-6.58e-09	4.20e+01	1.60e-03
56491.56	56493.90	2.34e+00	10	2.58e-09	6.70e+01	6.10e-04
56590.15	56592.52	2.37e+00	10	9.51e-09	3.49e+01	2.31e-03
56691.13	56693.47	2.34e+00	10	1.53e-08	2.75e+01	3.63e-03
56793.95	56796.30	2.35e+00	10	8.50e-09	3.69e+01	2.03e-03
56892.56	56894.92	2.36e+00	10	1.01e-09	1.07e+02	2.44e-04
56991.23	56993.52	2.29e+00	10	-3.85e-10	1.73e+02	8.72e-05
57092.19	57094.51	2.32e+00	10	-9.85e-09	3.43e+01	2.29e-03
57190.76	57193.14	2.38e+00	10	-9.49e-09	3.49e+01	2.32e-03
57291.76	57294.11	2.35e+00	10	2.98e-09	6.23e+01	7.09e-04
57390.34	57392.72	2.38e+00	10	-1.09e-08	3.26e+01	2.66e-03
57493.18	57495.53	2.35e+00	10	1.09e-09	1.03e+02	2.60e-04
57591.81	57594.14	2.33e+00	10	2.86e-09	6.36e+01	6.73e-04
57690.42	57692.76	2.34e+00	10	6.52e-10	1.33e+02	1.54e-04
57791.36	57793.74	2.38e+00	10	-2.50e-09	6.81e+01	6.10e-04
57890.00	57892.35	2.35e+00	10	-3.17e-09	6.04e+01	7.58e-04
57990.99	57993.33	2.34e+00	10	7.07e-09	4.05e+01	1.67e-03
58093.80	58096.16	2.36e+00	10	-7.42e-09	3.95e+01	1.79e-03
58192.39	58193.19	8.01e-01	20	-3.26e-09	5.96e+01	9.02e-05
58291.04	58293.38	2.34e+00	10	-5.01e-09	4.80e+01	1.19e-03
58392.00	58394.36	2.36e+00	10	-5.35e-09	4.65e+01	1.29e-03
58490.58	58492.93	2.35e+00	10	-1.41e-08	2.86e+01	3.37e-03
58591.61	58593.94	2.33e+00	10	6.70e-09	4.16e+01	1.57e-03
58690.19	58692.53	2.34e+00	10	-9.98e-09	3.41e+01	2.36e-03
58793.00	58795.38	2.38e+00	10	4.58e-10	1.59e+02	1.12e-04

Table 5.3: Time of observations, spin change rate and decoherence time scale of OAO 1657-415

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
54694.10	54697.00	2.90	20	-1.04e-07	1.06e+01	3.76e-02
54791.03	54793.94	2.91	20	-1.01e-07	1.07e+01	3.70e-02
54892.60	54895.49	2.89	20	-1.82e-07	7.97e+00	6.56e-02
55010.44	55013.31	2.87	20	-1.80e-07	8.02e+00	6.39e-02
55091.11	55093.98	2.87	20	-1.95e-07	7.70e+00	6.96e-02
55192.68	55195.56	2.88	10	7.15e-07	4.02e+00	2.56e-01
55292.49	55295.40	2.91	10	6.50e-07	4.22e+00	2.38e-01
55391.17	55394.05	2.88	10	-8.62e-08	1.16e+01	3.09e-02
55491.01	55493.89	2.88	10	-2.15e-08	2.32e+01	7.70e-03
55592.57	55595.43	2.86	20	-1.22e-07	9.75e+00	4.30e-02
55699.92	55702.79	2.87	20	1.55e-08	2.74e+01	5.49e-03
55791.03	55793.92	2.89	10	4.88e-07	4.87e+00	1.76e-01
55917.00	55919.31	2.31	20	1.72e-07	8.20e+00	3.97e-02
56002.89	56005.78	2.89	30	-1.27e-07	9.56e+00	4.57e-02
56091.14	56093.99	2.85	10	9.14e-08	1.13e+01	3.20e-02
56198.47	56201.38	2.91	10	-3.89e-07	5.46e+00	1.42e-01
56302.95	56305.84	2.89	20	-4.91e-08	1.53e+01	1.77e-02
56394.07	56396.95	2.88	20	-3.84e-08	1.74e+01	1.38e-02
56490.99	56493.87	2.88	20	1.97e-07	7.67e+00	7.06e-02
56592.54	56595.42	2.88	10	5.12e-07	4.75e+00	1.84e-01
56723.74	56726.63	2.89	30	-7.92e-08	1.21e+01	2.86e-02
56793.94	56796.84	2.90	20	-2.00e-07	7.60e+00	7.26e-02
56901.33	56904.21	2.88	50	-1.68e-07	8.31e+00	6.02e-02
56992.44	56995.34	2.90	10	4.65e-07	4.99e+00	1.69e-01
57096.91	57099.79	2.88	10	-1.43e-07	9.01e+00	5.11e-02
57190.91	57193.82	2.91	10	4.23e-07	5.23e+00	1.55e-01
57305.83	57308.74	2.91	30	-1.39e-07	9.13e+00	5.08e-02
57399.88	57402.76	2.88	20	-2.08e-07	7.47e+00	7.46e-02
57490.96	57493.85	2.89	20	3.37e-07	5.86e+00	1.22e-01
57590.81	57593.71	2.90	20	-8.70e-08	1.15e+01	3.17e-02
57692.37	57695.27	2.90	10	5.29e-07	4.68e+00	1.92e-01
57793.93	57796.81	2.88	20	-3.92e-07	5.44e+00	1.40e-01

Table 5.3: Time of observations, spin change rate and decoherence time scale of OAO 1657-415 (continued)

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end}-T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
57890.85	57893.75	2.90	10	-5.99e-07	4.40e+00	2.17e-01
57992.46	57995.32	2.86	10	-4.94e-08	1.53e+01	1.75e-02
58099.79	58102.70	2.91	20	9.30e-08	1.12e+01	3.40e-02
58214.69	58217.60	2.91	30	-1.10e-08	3.24e+01	4.04e-03
58298.27	58301.15	2.88	20	-1.44e-07	8.96e+00	5.17e-02
58392.29	58395.20	2.91	20	-1.25e-07	9.62e+00	4.57e-02
58499.68	58502.58	2.90	30	-3.97e-09	5.40e+01	1.44e-03
58598.35	58601.23	2.88	30	1.52e-07	8.72e+00	5.45e-02
58692.34	58695.24	2.90	10	-1.74e-07	8.17e+00	6.30e-02
58799.72	58802.63	2.91	20	-1.67e-07	8.32e+00	6.12e-02

Table 5.4: Time of observations, spin change rate and decoherence time scale of GX 301-2

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end}-T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
54692.08	54693.96	1.88	20	2.16e-08	2.31e+01	3.31e-03
54792.00	54793.97	1.97	10	-1.81e-08	2.53e+01	3.03e-03
54890.00	54891.97	1.97	10	3.81e-10	1.74e+02	6.38e-05
54990.02	54991.96	1.94	10	2.18e-08	2.31e+01	3.55e-03
55090.02	55091.99	1.97	10	3.78e-08	1.75e+01	6.33e-03
55190.02	55191.99	1.97	10	-1.90e-08	2.47e+01	3.19e-03
55290.01	55291.81	1.80	10	9.91e-09	3.42e+01	1.38e-03
55390.09	55391.98	1.89	10	2.56e-07	6.73e+00	3.95e-02
55492.00	55494.00	2.00	10	-1.73e-08	2.59e+01	2.99e-03
55590.00	55592.00	2.00	10	7.95e-09	3.82e+01	1.37e-03
55690.07	55691.97	1.90	10	-4.19e-08	1.66e+01	6.52e-03
55790.02	55791.97	1.95	10	2.56e-08	2.13e+01	4.20e-03

Table 5.4: Time of observations, spin change rate and decoherence time scale of GX 301-2 (continued)

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	No. of days on each side of T to estimate $\dot{\nu}$	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
55890.02	55892.00	1.98	10	-1.60e-08	2.69e+01	2.71e-03
55990.02	55991.96	1.94	10	-5.78e-08	1.42e+01	9.40e-03
56092.02	56093.99	1.97	10	-8.12e-09	3.77e+01	1.36e-03
56190.01	56191.96	1.95	10	-3.21e-08	1.90e+01	5.28e-03
56290.01	56291.97	1.96	10	-5.76e-08	1.42e+01	9.54e-03
56390.20	56391.90	1.70	10	-4.32e-08	1.64e+01	5.39e-03
56490.04	56491.99	1.95	10	-4.50e-08	1.60e+01	7.38e-03
56590.08	56591.89	1.81	10	7.89e-09	3.83e+01	1.12e-03
56692.13	56694.00	1.87	10	-2.57e-08	2.12e+01	3.88e-03
56790.00	56792.00	2.00	10	-1.35e-08	2.93e+01	2.34e-03
56890.11	56891.99	1.88	10	-5.95e-08	1.39e+01	9.07e-03
56990.44	56991.99	1.55	10	4.47e-09	5.09e+01	4.62e-04
57090.34	57091.89	1.55	10	1.54e-08	2.74e+01	1.60e-03
57200.05	57201.99	1.94	20	-6.33e-09	4.28e+01	1.03e-03
57290.05	57292.00	1.95	10	4.22e-08	1.66e+01	6.93e-03
57390.03	57391.98	1.95	10	-2.78e-08	2.04e+01	4.56e-03
57490.01	57491.98	1.97	10	1.03e-08	3.36e+01	1.72e-03
57590.25	57591.90	1.65	10	5.35e-09	4.65e+01	6.28e-04
57692.01	57693.97	1.96	10	-1.56e-08	2.72e+01	2.58e-03
57794.00	57796.00	2.00	10	4.33e-08	1.63e+01	7.49e-03
57890.00	57891.93	1.93	10	-4.15e-08	1.67e+01	6.68e-03
57990.01	57992.00	1.99	20	9.20e-09	3.55e+01	1.57e-03
58090.00	58092.00	2.00	10	3.47e-08	1.83e+01	5.99e-03
58190.02	58191.98	1.96	20	-1.87e-08	2.49e+01	3.11e-03
58290.13	58291.99	1.86	10	2.36e-08	2.22e+01	3.52e-03
58392.29	58395.20	2.91	10	-2.65e-08	2.09e+01	9.70e-03
58499.68	58502.58	2.90	10	5.40e-07	4.63e+00	1.96e-01
58598.35	58601.23	2.88	10	-6.88e-08	1.30e+01	2.46e-02
58692.34	58695.24	2.90	10	-7.19e-08	1.27e+01	2.61e-02
58799.72	58802.63	2.91	10	1.99e-08	2.41e+01	7.28e-03

Table 5.5: Time of observations, spin change rate and decoherence time scale of GX 301-2 during the first rapid spin-up episode

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
54821.02	54822.00	0.98	9.87e-08	1.08e+01	4.09e-03
54822.01	54823.99	1.98	7.38e-08	1.25e+01	1.25e-02
54824.00	54825.98	1.98	4.08e-08	1.68e+01	6.91e-03
54826.00	54827.97	1.97	-9.50e-09	3.49e+01	1.59e-03
54828.00	54830.00	2.00	-2.40e-08	2.20e+01	4.15e-03
54830.00	54832.00	2.00	8.86e-08	1.14e+01	1.53e-02
54834.03	54836.00	1.97	3.56e-07	5.70e+00	5.97e-02
54836.02	54837.99	1.97	3.85e-07	5.48e+00	6.45e-02
54838.02	54839.98	1.96	3.26e-07	5.96e+00	5.41e-02
54840.01	54841.97	1.96	2.59e-07	6.68e+00	4.30e-02
54842.00	54843.97	1.97	1.99e-07	7.62e+00	3.34e-02
54844.00	54846.00	2.00	2.18e-07	7.29e+00	3.77e-02
54846.07	54848.00	1.93	2.26e-07	7.15e+00	3.64e-02
54856.02	54857.97	1.95	-1.72e-07	8.19e+00	2.83e-02
54858.01	54859.96	1.95	-1.76e-07	8.11e+00	2.89e-02
54860.00	54861.95	1.95	-8.69e-08	1.15e+01	1.43e-02
54862.06	54863.94	1.88	-7.03e-08	1.28e+01	1.07e-02
54864.05	54866.00	1.95	-1.13e-07	1.01e+01	1.85e-02
54866.11	54867.99	1.88	2.56e-08	2.13e+01	3.91e-03

Table 5.6: Time of observations, spin change rate and decoherence time scale of GX 301-2 during the second rapid spin-up episode

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
55360.01	55361.96	1.95	-2.20e-08	2.29e+01	3.62e-03
55362.00	55363.96	1.96	-3.30e-08	1.87e+01	5.47e-03
55370.05	55371.82	1.77	1.29e-07	9.49e+00	1.74e-02
55372.33	55373.81	1.48	3.56e-07	5.70e+00	3.37e-02
55374.04	55375.98	1.94	4.51e-07	5.07e+00	7.33e-02
55376.03	55377.97	1.94	4.65e-07	4.99e+00	7.56e-02
55378.02	55379.72	1.70	5.24e-07	4.70e+00	6.54e-02
55380.01	55381.71	1.70	3.95e-07	5.41e+00	4.93e-02
55382.00	55383.94	1.94	2.55e-07	6.73e+00	4.15e-02
55384.06	55386.00	1.94	1.92e-07	7.75e+00	3.13e-02
55386.05	55387.99	1.94	2.31e-07	7.07e+00	3.76e-02
55388.04	55389.98	1.94	2.52e-07	6.78e+00	4.10e-02
55390.09	55391.98	1.89	3.24e-07	5.98e+00	5.00e-02
55392.08	55393.97	1.89	3.47e-07	5.78e+00	5.35e-02
55394.01	55395.96	1.95	2.86e-07	6.36e+00	4.70e-02
55396.06	55397.96	1.90	2.44e-07	6.88e+00	3.81e-02
55398.05	55400.00	1.95	1.06e-07	1.05e+01	1.74e-02
55400.04	55402.00	1.96	-7.05e-08	1.28e+01	1.17e-02
55402.04	55404.00	1.96	-5.49e-08	1.45e+01	9.12e-03
55404.00	55405.99	1.99	-1.58e-08	2.71e+01	2.70e-03
55406.00	55407.98	1.98	3.32e-08	1.87e+01	5.62e-03
55408.00	55410.00	2.00	-2.93e-08	1.99e+01	5.06e-03

Table 5.7: Time of observations, spin change rate and decoherence time scale of GX 301-2 during the third rapid spin-up episode

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
58458.02	58459.96	1.94	8.24e-08	1.19e+01	1.34e-02
58460.06	58462.00	1.94	1.21e-07	9.80e+00	1.96e-02
58462.03	58463.99	1.96	2.78e-08	2.04e+01	4.61e-03
58464.01	58465.97	1.96	-9.52e-08	1.10e+01	1.58e-02
58466.00	58468.00	2.00	-1.38e-07	9.15e+00	2.39e-02
58468.00	58470.00	2.00	-1.79e-07	8.04e+00	3.09e-02
58486.03	58488.00	1.97	5.33e-07	4.66e+00	8.93e-02
58488.01	58489.98	1.97	5.33e-07	4.66e+00	8.93e-02
58490.00	58492.00	2.00	4.75e-07	4.94e+00	8.21e-02
58492.01	58493.98	1.97	4.60e-07	5.01e+00	7.72e-02
58494.05	58496.00	1.95	5.03e-07	4.80e+00	8.26e-02
58496.01	58497.99	1.98	3.83e-07	5.49e+00	6.49e-02
58498.00	58500.00	2.00	4.67e-07	4.98e+00	8.07e-02
58500.02	58501.98	1.96	6.00e-07	4.39e+00	9.96e-02
58502.03	58503.99	1.96	8.14e-07	3.77e+00	1.35e-01
58504.02	58505.98	1.96	7.48e-07	3.93e+00	1.24e-01
58506.00	58508.00	2.00	5.70e-07	4.51e+00	9.85e-02
58508.01	58509.99	1.98	4.64e-07	5.00e+00	7.85e-02
58510.03	58511.99	1.96	4.23e-07	5.23e+00	7.02e-02
58530.33	58531.94	1.61	1.25e-07	9.62e+00	1.40e-02
58532.25	58533.99	1.74	1.68e-07	8.29e+00	2.20e-02
58534.01	58535.97	1.96	1.69e-07	8.26e+00	2.81e-02
58536.02	58537.99	1.97	1.72e-07	8.20e+00	2.89e-02
58538.00	58539.97	1.97	1.75e-07	8.13e+00	2.93e-02
58540.02	58541.99	1.97	1.60e-07	8.51e+00	2.68e-02
58542.00	58544.00	2.00	2.19e-07	7.26e+00	3.79e-02
58544.00	58545.99	1.99	2.86e-07	6.36e+00	4.90e-02
58546.00	58547.99	1.99	3.32e-07	5.91e+00	5.68e-02
58566.23	58567.85	1.62	-1.77e-07	8.08e+00	2.01e-02
58568.21	58569.83	1.62	-8.91e-08	1.14e+01	1.01e-02

Table 5.7: Time of observations, spin change rate and decoherence time scale of GX 301-2 during the third rapid spin-up episode (continued)

T_{start} (MJD)	T_{end} (MJD)	$T = T_{end} - T_{start}$ (day)	$\dot{\nu}$ (Hz/day)	$\sqrt{(1/ \dot{\nu})}$ (day)	$\frac{1}{2} \dot{\nu} T^2$
58570.19	58571.81	1.62	-1.12e-07	1.01e+01	1.27e-02
58572.17	58573.86	1.69	1.06e-07	1.05e+01	1.30e-02
58574.16	58575.91	1.75	1.12e-07	1.02e+01	1.48e-02
58576.14	58577.90	1.76	3.56e-09	5.70e+01	4.76e-04
58578.05	58579.95	1.90	8.54e-08	1.16e+01	1.33e-02
58580.04	58582.00	1.96	5.02e-08	1.52e+01	8.33e-03
58582.02	58583.98	1.96	4.63e-08	1.58e+01	7.68e-03
58584.03	58586.00	1.97	5.55e-08	1.44e+01	9.30e-03
58586.01	58587.95	1.94	4.50e-08	1.60e+01	7.32e-03
58588.02	58589.99	1.97	-7.68e-08	1.23e+01	1.29e-02
58590.00	58591.98	1.98	-8.29e-08	1.18e+01	1.40e-02
58592.02	58593.99	1.97	-2.62e-08	2.10e+01	4.39e-03
58594.00	58595.95	1.95	-2.22e-07	7.22e+00	3.65e-02
58596.02	58598.00	1.98	-2.53e-07	6.76e+00	4.29e-02

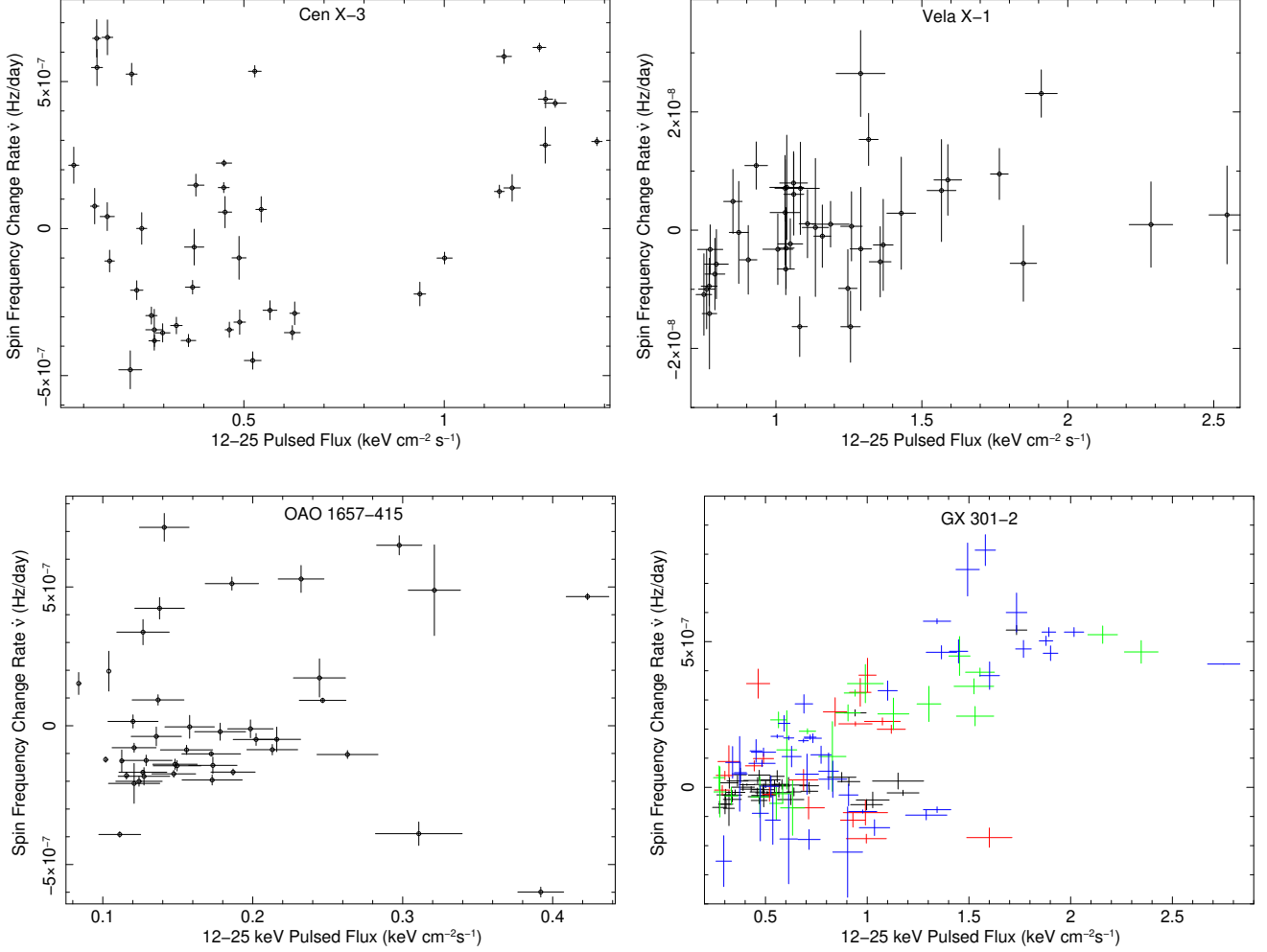


Figure 5.3: These four figures shows the spin change rates with the pulsed flux measurements for Cen X-3 (top left), Vela X-1 (top right), OAO 1657-415 (bottom left) and GX 301-2 (bottom right). In the bottom right panel, the coloured data points are for the three rapid spin-up episodes of GX 301-2. The red, green and blue data points are for the spin-up episodes around MJD 54825-54860, MJD 55367-55403 and MJD 58465-58567 respectively.

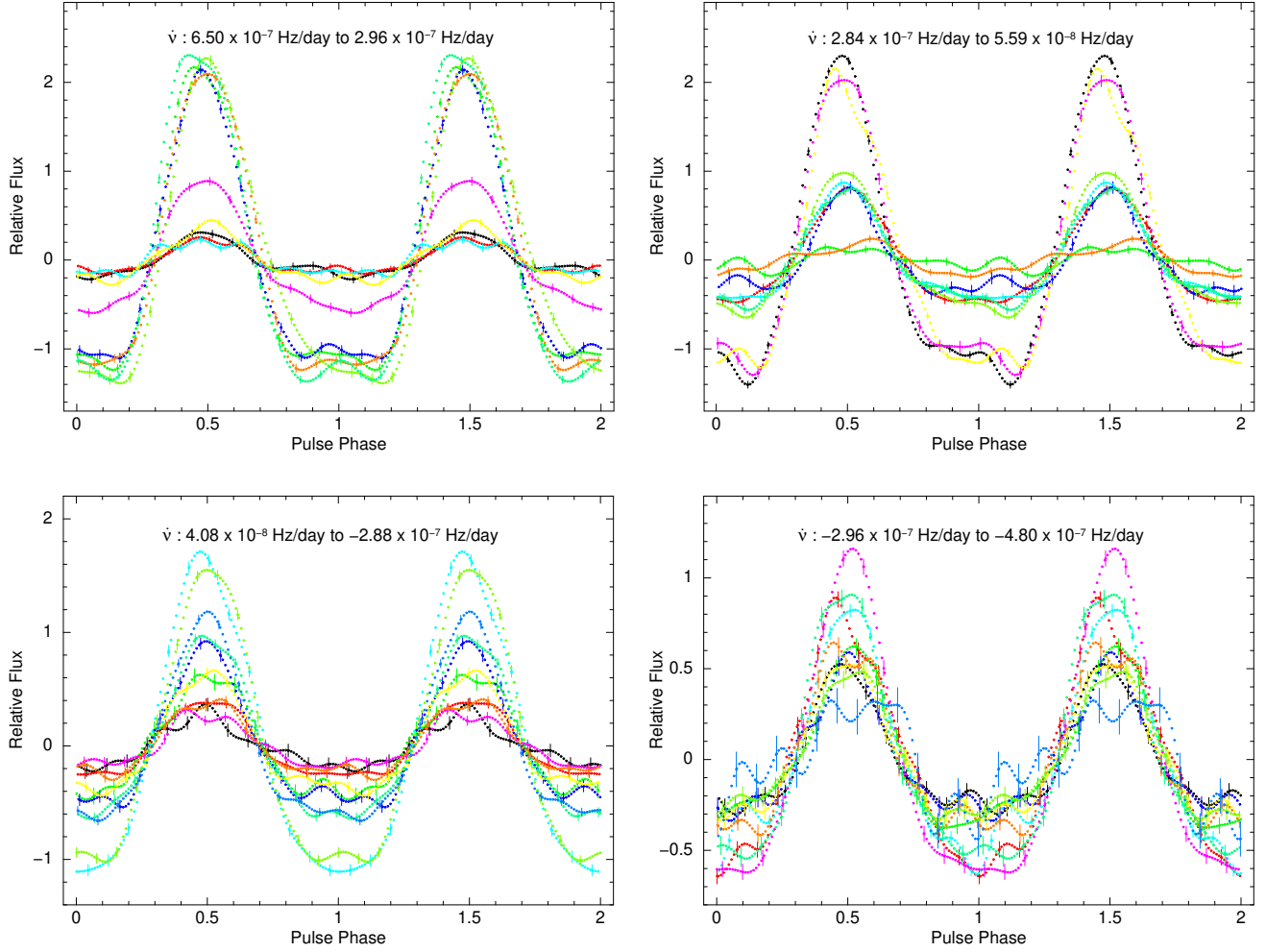


Figure 5.4: These four figures show the pulse profiles of Cen X-3 in four different range of spin change rates.

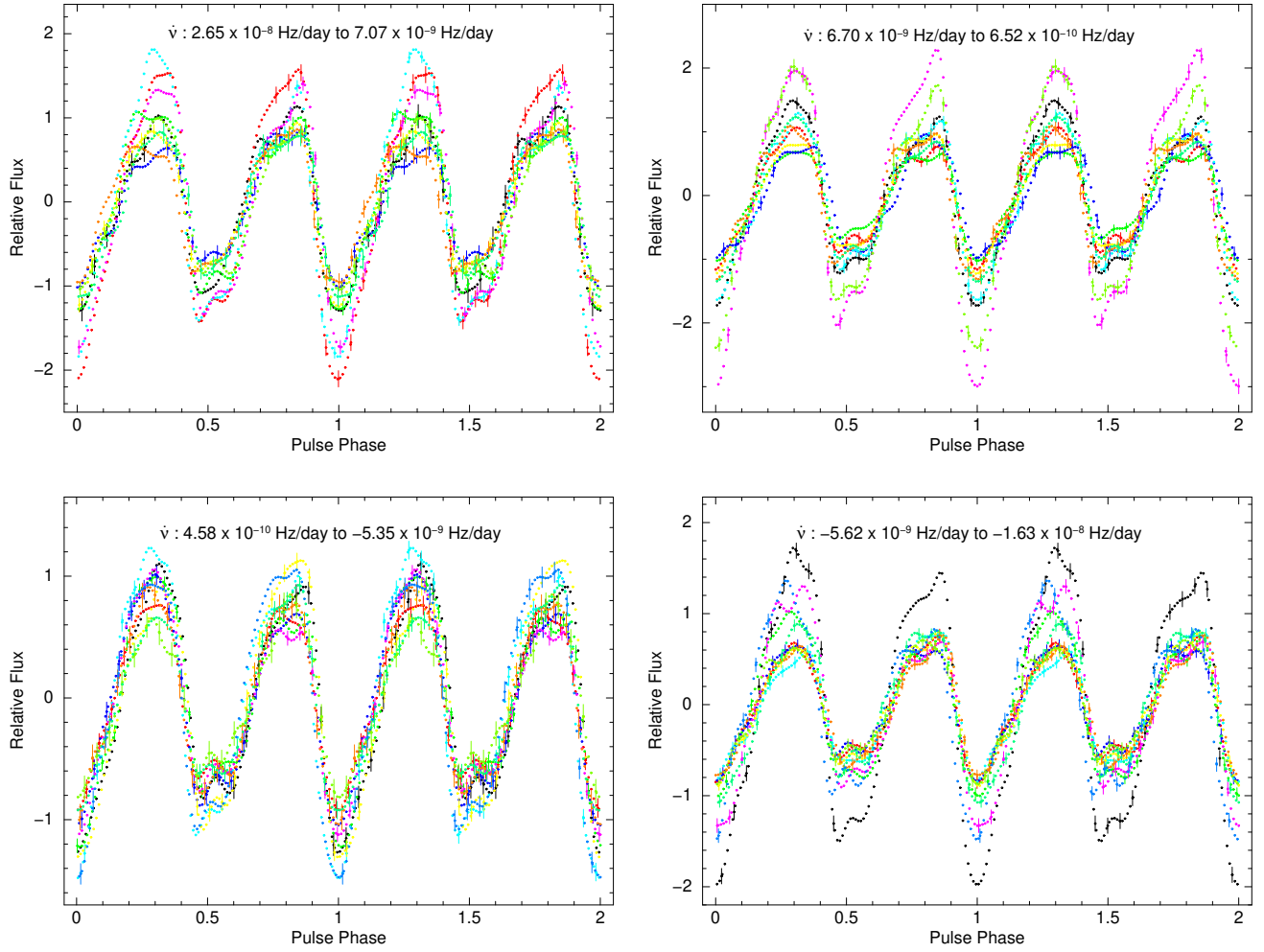


Figure 5.5: These four figures show the pulse profiles of Vela X-1 in four different range of spin change rates.

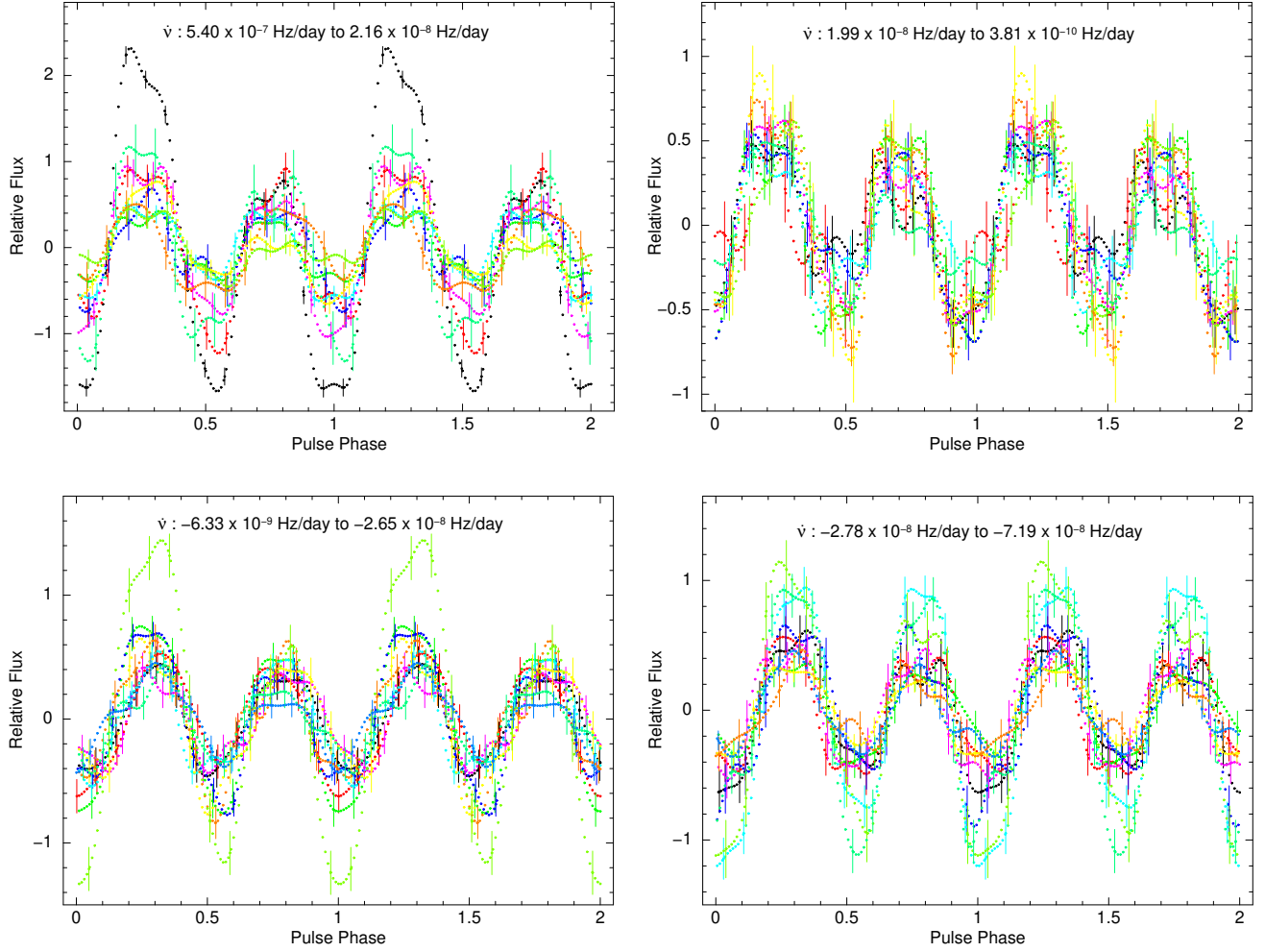


Figure 5.6: These four figures show the pulse profiles of GX 301-2 in four different range of spin change rates when the source goes through random spin-up and down.

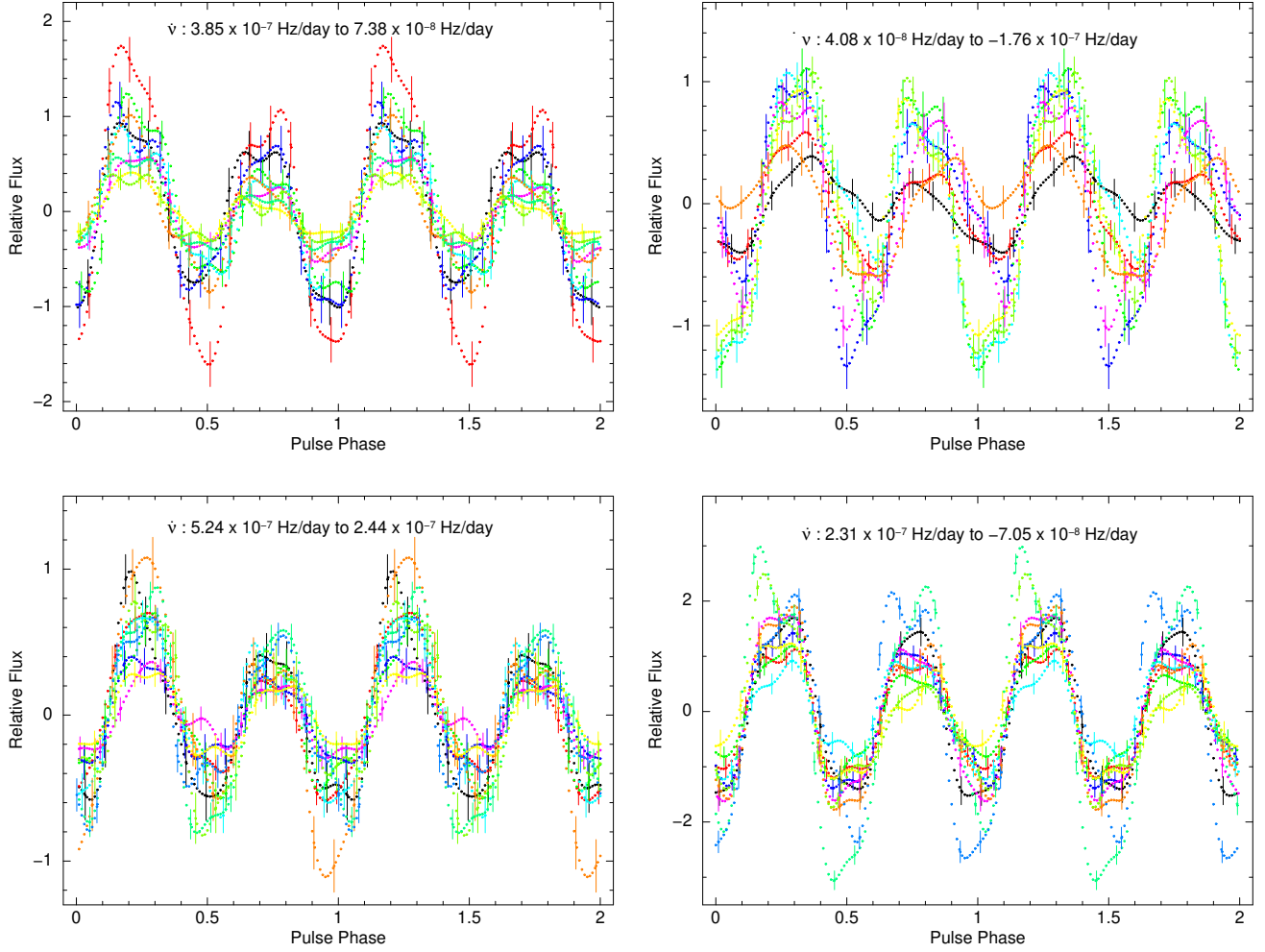


Figure 5.7: These four figures show the pulse profiles during the first (top) and second (bottom) rapid spin-up episode of GX 301-2.

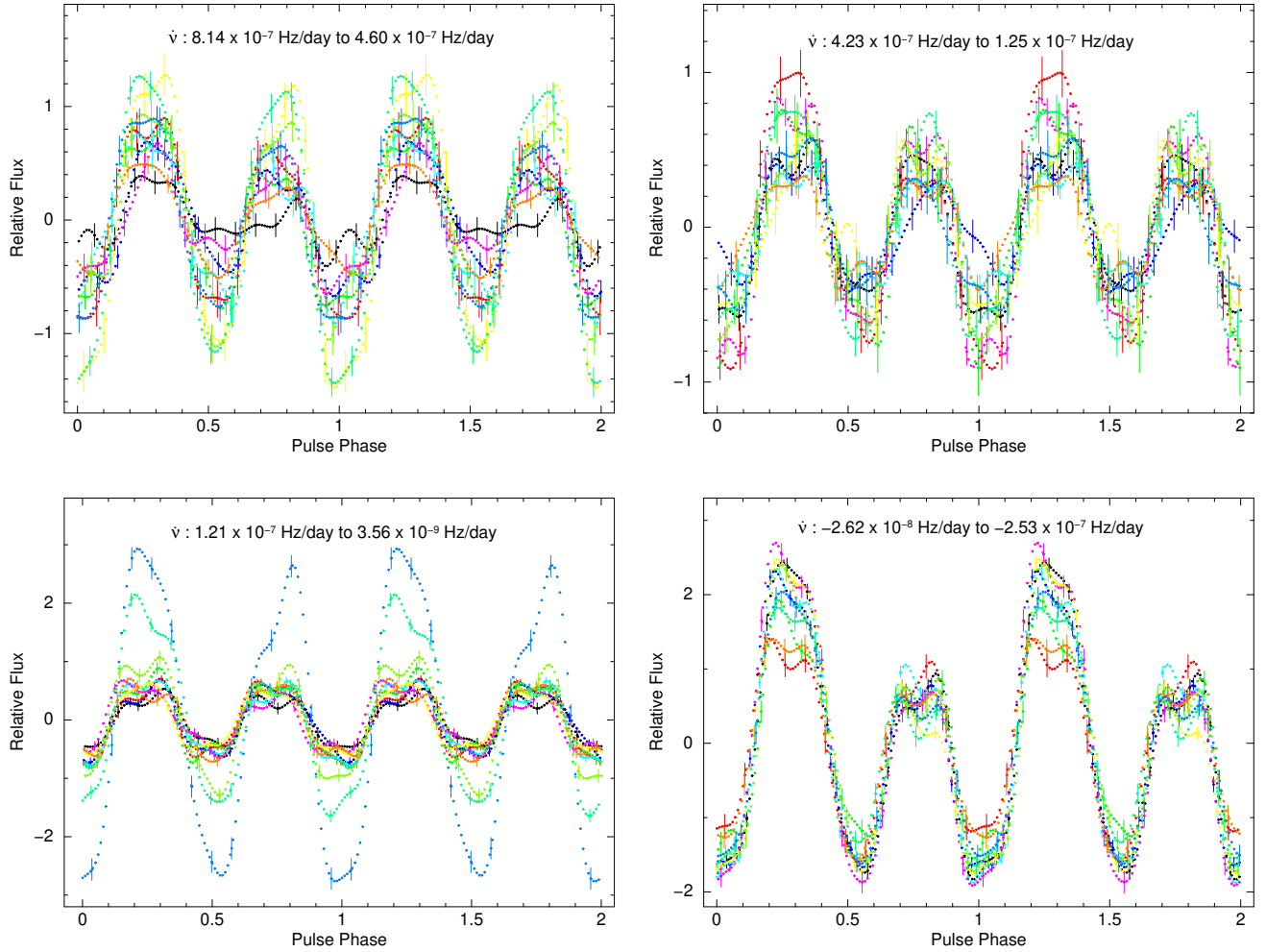


Figure 5.8: These four figures show the pulse profiles during the third rapid spin-up episode of GX 301-2.

5.3 Results and discussion

The HMXB pulsars under this study are all highly magnetized neutron stars. In the presence of a strong magnetic field the accreted matter is channelled along the magnetic field lines forming two polar caps. The mass accretion rate and the local hydro-dynamical changes leads to the formation of the X-ray emission region and leads to distinctive beaming patterns. The emitted flux from the polar caps from the rotating neutron star is detected as pulses. The pulse profiles of the accreting pulsars are complex in nature. They vary in shape and amplitude and they are the unique characteristics of the pulsars. Despite of multiple theoretical efforts, the morphology of the pulse profiles is not clearly understood till date. However, the pulse profiles show strong correlation with the X-ray luminosity and therefore the mass accretion rate. Also, the profiles changes shape from from one energy band to another. In the low energy band dip features appear in the profiles as a result of the absorption by the phase locked accretion streams and disappear mostly above 7-15 keV (Devasia et al. 2011, Maitra et al. 2012). There are four major factors which constitute the pulse profile for a particular pulsar. They are : 1) The alignment of the spin and magnetic dipole axes and the angle they create with the observer's line of sight (Harding and Muslimov 1998), 2) the beaming pattern of the emitted flux, 3) the deflection of the emitted flux by the relativistic effects due to the high gravitational field of the neutron star (Riffert and Meszaros 1988, Kraus 2001), and 4) the configuration of the magnetic field which may not be symmetric dipole configuration (Kraus et al. 1996; 1995). For a particular pulsar, the change in the pulse shape is most likely to occur due to the change in the beaming pattern.

In this work, we have looked at the phase aligned pulse profiles of three pulsars. All this profiles were extracted from the 12-25 keV energy band. The statistics of the data was better than that of the other two bands (8-12 keV and 25-50 keV). Another advantage is that the 12-25 keV is a sufficiently high energy range where the pulse shapes are not much affected by the absorption related modulations. The overall shape of the pulses remains unchanged for Cen X-3. They are single peaked and the beaming pattern remains the same. The spin frequency change rate is a direct measurement of the transfer of angular momentum transferred to the neutron star. Even though Cen X-3 is disk fed, the transfer of angular momentum by the accreted matter is not well understood for this source. The same is true for Vela X-1 and GX 301-2. They have double peaked pulse profiles. It is possibly due to the pencil beam emission from both the magnetic poles and

emission from both poles are visible to the observer. Double peaked pulse profiles are quite common in many other pulsars like XTE J1946+274 (Maitra and Paul 2013), 4U 1538-52 (Bildsten et al. 1997). The erratic spin up and down behaviour of these persistent pulsars can be a result of the continuous change in the flow pattern of the accreted matter onto the neutron star, so that there is no net transfer of angular momentum. Therefore, the pulsar spins up and down in smaller time scales keeping the rotation speed almost unchanged. There were several theoretical attempts to establish this type of unstable accretion flow with alternating sign of angular momentum (for example Fryxell and Taam 1988, Matsuda et al. 1991). Our aim was to see if this unstable accretion flow incorporates any significant change to the geometry of the emitted radiation or the beaming pattern. A clear correlation between the pulse shapes and spin change rate could clearly indicate the continuous change in the accretion flow and the flow pattern for the transfer of positive or negative angular momentum. However, we do not see any such correlation. One possibility is that the unstable flow of accreted matter does not change the overall emission pattern of the pulsar. The emission pattern solely depends on the accretion rate. Another thing is that the pulse profiles detected by *Fermi* have large error bars associated with them which prohibits from achieving a smooth and accurate shape. Multiple pointing observations with several intervals with adequate (depending on the pulse period) time resolution over a long period of time can help us to probe this further.

Chapter 6

Summary and future work



Summary

In this thesis, we have investigated X-ray variabilities at different time-scales and probed certain aspects of individual high mass X-ray binary systems. We carried out timing and spectroscopic studies to understand the X-ray emission mechanisms and the physics of accretion. We have summarized the results of our studies here.

- In chapter 3, we have studied a unique partial eclipse seen in IGR J16393-4643. The orbital profile of the High Mass X-ray binary IGR J16393-4643 shows a dip in its X-ray intensity, which was previously interpreted as an eclipse. Unlike most eclipsing HMXBs, where the X-ray eclipses are about two orders of magnitude fainter compared to the out of eclipse emission, this particular eclipse like feature is narrow and partial, casting doubt if it is indeed an eclipse.

We analyzed a large number of observations with *Swift*-XRT, covering the entire orbital phase. The soft X-ray observations also show this low-intensity phase, which is about 30% of the intensity during the rest of the orbit. We carried out orbital phase-resolved spectroscopy to compare the change in the spectral parameters inside and outside of this low-intensity state. Our results indicate that this low-intensity state might not be an eclipse, as previously thought, but absorption in the stellar corona. We have also provided constraints on the inclination angle of the binary for grazing eclipse caused by the stellar corona.

- The Be X-ray binary GRO J2058+42 went through its second Type-II outburst during March-April 2019, lasting for about 50 days. This outburst was detected with all the operating all-sky X-ray monitors like the *Fermi*-GBM, *Swift*-BAT, and *MAXI*-GSC. Two TOO (target of opportunity) observations by *NuSTAR* were also made, one during the rise and the other during the decay of the outburst. After the outburst, the binary went to a low-luminosity state and was observed again by *NuSTAR*.

In chapter 4, we have analyzed the broadband characteristics of the pulsar for the first time. We found that the pulse profiles are strongly energy dependent both during the outburst and the low luminosity state. We determined the spin period of the pulsar from each of the *NuSTAR* observations.

During the outburst, we noticed at least four different pulse components at low energy (< 20 keV), which evolves to a single-peaked profile at high energy (> 30 keV).

In each of the narrow energy bands, the pulse profiles are nearly identical in the two *NuSTAR* observations during the outburst. The spectra from both the observations are fitted well to a power-law with a Fermi-Dirac type high energy cutoff. We ruled out the presence of a cyclotron line in the pulse phase averaged X-ray spectrum in the *NuSTAR* band with an optical depth greater than 0.15. An iron emission line has been detected in both the *NuSTAR* spectra with an equivalent width of about 125 eV.

During the low-luminosity state, the pulse profiles have three peaks in the lowest energy band (3-6 keV) and it becomes double peaked above 20 keV. The spectra (FPMA and FPMB) were best fitted with an absorbed power law with HIGHECUT. We do not notice any emission or absorption line in the spectra. The observed luminosity is higher than the typical quiescence luminosity of other Be X-ray binaries in quiescence. The luminosity and the power-law shape of the spectra point towards a non-thermal X-ray emission. It indicates that the source might be accreting from a cold disk.

We also looked at the dependence of the spin-up rate on the luminosity during the outburst using the *Fermi*-GBM data and estimated the magnetic field strength from that, which came out to be much higher compared to other known BeXRB pulsars. Lastly, We have discussed the inadequacy of the torque-luminosity relation for the determination of the magnetic field strength of neutron stars.

- The four persistent pulsars Cen X-3, GX 301-2, Vela X-1 and OAO 1657-415 go through rapid spin-up and spin-down episodes. The long term X-ray flux history and the spin frequency history of these pulsars indicate that mass accretion rate cannot be the only factor behind the transfer of the torque. Such torque behaviour can be attributed to the unstable accretion flow that transfers both positive and negative angular momentum to the neutron star. The shape of the pulse profiles depends on the geometry of the accretion column and the polar hot spot, and hence it is a key to observe a change in accretion mode. In chapter 6, we have investigated the shape of the pulse profiles to look for any possible changes in the accretion geometry at different accretion torque to understand the torque behaviour of these pulsars.

We were provided with the pulse profiles of these pulsars from the GBM team, at different time windows that are evenly distributed over the past decade. We deter-

mined the spin change rates during those time windows. The spin change rates do not show any obvious correlation with the X-ray luminosity. We stacked the pulse profiles of Cen X-3, Vela X-1 and GX 301-2 with respect to their decreasing order of spin change rates. We did not continue our study with OAO 1657-415 because of the large contribution of decoherence.

From our study, we do not see any correlation between the shape of the pulse profiles and the local spin change rates for any of the pulsars. This indicates that the transferred angular momentum by the unstable accretion flow in Cen X-3, Vela X-1 and GX 301-2, do not change the overall emission pattern.

Future Work

The work carried out in this thesis opens up possibility of further investigations of our chosen HMXBs. Our future plans are enumerated as follows:

- The study of the highly absorbed HMXB IGR J16393-4643 was carried out by systematic analysis of the archived *Swift*-XRT data. It was the only available data that covered the partial eclipse phase. The *Swift*-XRT observations had small exposure times with limited statistics. Therefore, we could not put a constraint on the spectral parameters like absorption column density, photon index etc. Although, our study indicates that the partial eclipse phase might occur due to grazing eclipse caused by the stellar corona, it requires further probing to establish the exact nature of the partial eclipse phase. We have requested several proposals for new observations. The *AstroSAT* proposal had been accepted and we received the data. Guest observation proposals were also written for *XMM-Newton* observatory and *NuSTAR*. It will be essential to estimate the spectral parameters with higher accuracy. The estimated ratio of the spectral parameters (column density and equivalent width of iron $K\alpha$ line) inside and outside the eclipse phase will reveal the geometric configuration and the wind structure of this system.
- The Be X-ray binary pulsars can act as unique laboratories to understand the accretion mechanism of strongly magnetized neutron stars over a broad range of luminosity. In general, they manifest themselves through outburst activities, and then transit to the quiescence state with very low luminosity or no X-ray activity at

all. The nature of the X-ray emission in the low luminosity state varies from source to source and is governed by the spin period and the magnetic field strength of the neutron star. The detection of pulsations during the quiescence is an open question.

The low luminosity X-ray emission of GRO J2058+42 is likely to arise from the accretion from a cold disk. It shows pulsations, and the *NuSTAR* spectra are well described by a power-law suggesting a non-thermal origin of the observed emission. The cutoff power-law nature is in contrast with the spectral nature of few other BeXRB pulsars, which show clear pulsations, as their spectra during the quiescence are usually described by a two-component comptonization model (Doroshenko et al. 2012, Tsygankov et al. 2019). GRO J2058+42 makes an interesting candidate to better understand the disk-magnetospheric interactions by studying its low luminosity regime. We want to complement our results from the *NuSTAR* observation with the additional study of *NICER* observations. We will carry out the timing and spectroscopic analysis of the low luminosity state of this source to understand its behaviour in the soft X-ray band.

- The study to find possible relation between the pulse profiles and accretion torque behaviour of persistent pulsars is new and instrumental to carry forward such analysis with pointing observations with several intervals. Availability of observations with good time resolution at different accretion torques will help to map the pulse shapes with greater precision. The similarity or difference in the pulse shapes at different luminosity and torque for wind-fed and disk-fed regimes needs to be studied systematically. This type of analysis will be extremely useful for modelling the pulse profiles and advance our understanding on the accretion geometry of the X-ray binary pulsars.

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