

The Evolution of Compact Binary Star Systems

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Abstract

We review the formation and evolution of compact binary stars consisting of white dwarfs (WDs), neutron stars (NSs), and black holes (BHs). Binary NSs and BHs are thought to be the primary astrophysical sources of gravitational waves (GWs) within the frequency band of ground-based detectors, while compact binaries of WDs are important sources of GWs at lower frequencies to be covered by space interferometers (LISA). Major uncertainties in the current understanding of properties of NSs and BHs most relevant to the GW studies are discussed, including the treatment of the natal kicks which compact stellar remnants acquire during the core collapse of massive stars and the common envelope phase of binary evolution. We discuss the coalescence rates of binary NSs and BHs and prospects for their detections, the formation and evolution of binary WDs and their observational manifestations. Special attention is given to AM CVn-stars – compact binaries in which the Roche lobe is filled by another WD or a low-mass partially degenerate helium-star, as these stars are thought to be the best LISA verification binary GW sources.

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1 Introduction

Close binary stars consisting of two compact stellar remnants (white dwarfs (WDs), neutron stars (NSs), or black holes (BHs)) are considered as primary targets of the forthcoming field of gravitational wave (GW) astronomy since their orbital evolution is entirely controlled by emission of gravitational waves and leads to ultimate coalescence (merger) of the components. Close compact binaries can thus serve as testbeds for theories of gravity. The double NS(BH) mergers should be the brightest GW events in the 10–1000 Hz frequency band of the existing GW detectors like LIGO [16], VIRGO [5], or GEO 600 [346]. Such mergers can be accompanied by the release of a huge amount of electromagnetic energy in a burst and manifest themselves as short gamma-ray bursts (GRBs). Double WDs, especially interacting binary WDs observed as AM CVn-stars, are potential GW sources within the frequency band of the space GW interferometers like LISA [96] or future detectors [68]. The double WD mergers also stay among the primary candidate mechanisms for type Ia supernova (SN Ia) explosions, which are crucial in modern cosmological studies.

Compact binaries are the end products of the evolution of binary stars, and the main purpose of the present review is to describe the astrophysical knowledge on their formation and evolution. We shall discuss the present situation with the main parameters determining their evolution and the rates of coalescence of double NSs/BHs and WDs.

About 6% of the baryonic matter in the Universe is confined in stars [114]. The typical mass of a stationary star is close to the solar value $M_{\odot} \approx 2 \times 10^{33}$ g. The minimum mass of a stationary star at the main sequence (MS) is set by the condition of stable hydrogen burning in its core $M_{\min} \approx 0.08 M_{\odot}$ [203]. The maximum mass of solar composition stars inferred observationally is close to $150 M_{\odot}$ [104]; for very low metallicity stars it is derived by the linear analysis of pulsational stability and is close to $300 M_{\odot}$ [15]. Stars and stellar systems are formed due to the development of the gravitational (Jeans) instability in turbulized molecular clouds. The minimum protostellar mass is dictated by the opacity conditions in the collapsing fragments and is found to lie in the range 0.01 – $0.1 M_{\odot}$ in both analytical [343] and numerical calculations (see, e.g., [74]). It is established from observations that the mass distribution of main-sequence stars has a power-law shape [365, 255], $dN/dM \sim M^{-\beta}$, with $\beta = -1.2$ for $0.08 \lesssim M/M_{\odot} \lesssim 0.5$, $\beta = -2.2$ for $0.5 \lesssim M/M_{\odot} \lesssim 1.0$, and $\beta = -2.2$ to -3.2 for $1.0 \lesssim M/M_{\odot} \lesssim 150$ [199, 200].

The evolution of a single star is determined by its initial mass at the main sequence M_0 and the chemical composition. If $M_0 \lesssim 8$ – $12 M_{\odot}$, the carbon-oxygen (CO) (or oxygen-neon (ONe) at the upper end of the range) stellar core becomes degenerate and the evolution of the star ends up with the formation of a CO or ONe white dwarf. The formation of a WD is accompanied by the loss of stellar envelope by stellar wind in the red giant and asymptotic giant branch stages of evolution and ejection of a planetary nebula. The boundary between the masses of progenitors of WDs and NSs is not well defined and is, probably, between 8 and $12 M_{\odot}$ (cf. [161, 157, 348, 159, 115, 324, 349, 119, 377]).

At the upper boundary of the mass range of white dwarf progenitors, formation of ONe WDs is possible. The masses of stars that produce ONe WDs are still highly uncertain. However, strong observational evidence for their existence stems from the analysis of nova ejecta [405]. This variety of WDs is important in principle, because accretion induced collapse (AIC) of them may result in formation of neutron stars (see [292, 77] and references therein), but since for the purpose of detection of gravitational waves they are not different from the much more numerous CO-WDs, we will, as a rule, not consider them below as a special class.

If $M_0 \gtrsim (10$ – $12) M_{\odot}$, thermonuclear evolution proceeds until iron-peak elements are produced in the core. Iron cores are subjected to instabilities (neutronization, nuclei photodesintegration, or pair creation for the most massive stars) that lead to gravitational collapse. The core collapse of massive stars results in the formation of a neutron star or, for very massive stars, a black hole and is associated with the brightest astronomical phenomena such as supernova explosions

Initial mass [M_{\odot}]	remnant type	mean remnant mass [M_{\odot}]
$0.95 < M < 8-10$	WD	0.6
$8-10 < M < 25-30$	NS	1.35
$25-30 < M < 150$	BH	~ 10

Table 1: *Types of compact stellar remnants (the ranges of progenitor mass are shown for solar composition stars).*

(of type II, Ib, or Ib/c, according to the astronomical classification based on the spectra and light curves properties). If the pre-collapsing core retains significant rotation, powerful gamma-ray bursts lasting up to hundreds of seconds may be produced [456].

The boundaries between the masses of progenitors of WDs or NSs and NSs or BHs are fairly uncertain (especially for BHs). Typically accepted masses of stellar remnants for nonrotating solar chemical composition stars are summarized in Table 1.

For a more detailed introduction into the physics and evolution of stars the reader is referred to the classical textbook [67]. Formation and physics of compact objects is described in more detail in monographs [375, 35]. For a recent review of the evolution of massive stars and the mechanisms of core-collapse supernovae we refer to [457, 111, 197].

Most stars in the Galaxy are found in multiple systems, with single stars (including our own Sun) being rather exceptions than a rule (see for example [86, 128]). In the binary stars with sufficiently large orbital separations (“wide binaries”) the presence of the secondary component does not influence significantly the evolution of the components. In “close binaries” the evolutionary expansion of stars allows for a mass exchange between the components. In close binaries, the initial mass of the components at the zero-age main sequence (ZAMS) ceases to be the sole parameter determining their evolution. Consequently, the formation of compact remnants in binary stars differs from single stars. This is illustrated by Figure 1 which plots the type of the stellar remnant as a function of both initial mass and the radius of a star at the moment of the Roche-lobe overflow (RLOF). It is seen that wide binaries evolve as single stars, while for binaries with RLOF a new type of remnants appears – a helium WD, whose formation from a single star in the Hubble time is impossible¹.

Binaries with compact remnants are primary potential GW sources (see Figure 2). This figure plots the sensitivity of ground-based interferometer LIGO, as well as the space laser interferometer LISA, in terms of dimensionless GW strain h measured over 1 year. The strongest Galactic sources at all frequencies are the most compact double NSs and BHs. Double WDs (including AM CVn-stars) and ultra-compact X-ray binaries (NS + WD) appear to be promising LISA sources.

Double NS/BH systems result from the evolution of initially massive binaries, while double WDs are formed from the evolution of low-mass binaries. We shall consider them separately.

Binaries with NSs and BHs

Binary systems with components massive enough to produce NSs or BHs at the end of thermonuclear evolution may remain bound after two supernova explosions. Then, loss of energy and momentum by GWs controls entirely their evolution and gradual reduction of the binary separation may bring the components into contact. During the merger process $\sim 10^{52}$ erg are released as GWs [58, 59]. Such strong bursts of GWs can be reliably detected by the present-day ground-based GW detectors from distances up to several megaparsecs and are the most important targets for GW observatories such as LIGO, GEO, and VIRGO [122, 373].

¹The hydrogen burning time for single stars with $M_0 \lesssim 0.9 M_{\odot}$ exceeds the age of the Universe.

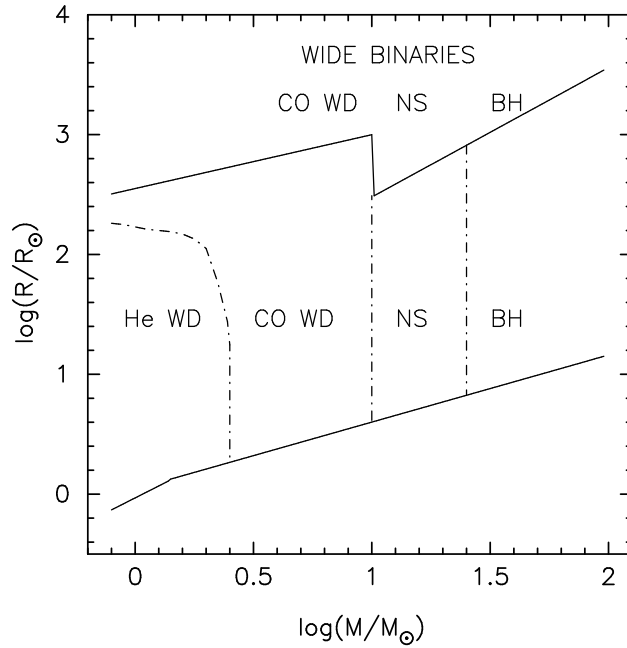


Figure 1: *Descendants of components of close binaries depending on the radius of the star at RLOF. The boundary between progenitors of He and CO-WDs is uncertain by several $0.1 M_{\odot}$, the boundary between WDs and NSs by $\sim 1 M_{\odot}$, while for the formation of BHs the lower mass limit may be even by $\sim 10 M_{\odot}$ higher than indicated.*

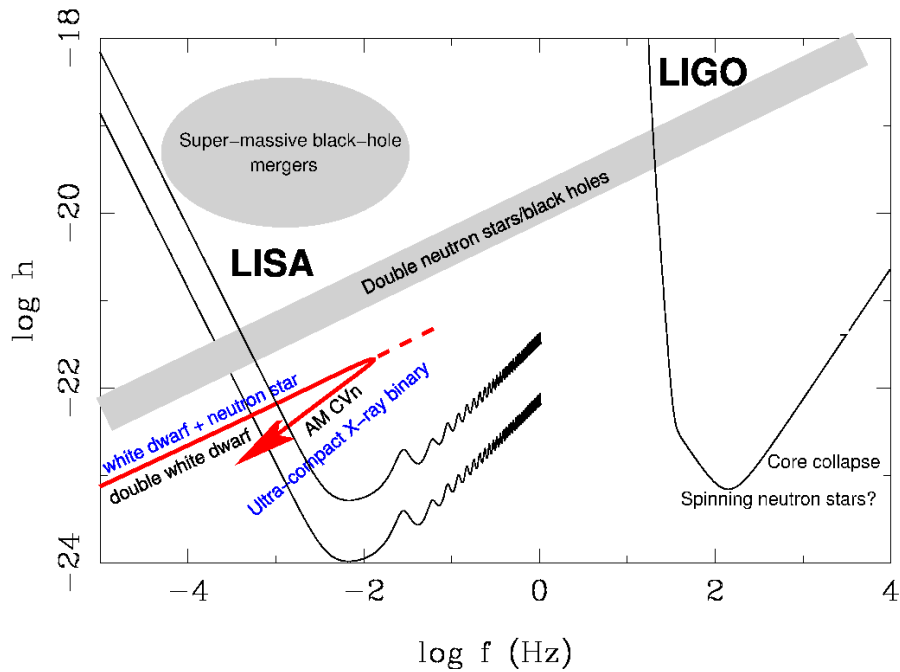


Figure 2: *Sensitivity limits of GW detectors and the regions of the $f-h$ diagram occupied by some of the potential GW sources. (Courtesy G. Nelemans.)*

The problem is to evaluate as accurately as possible (i) the physical parameters of the coalescing binaries (masses of the components and, if possible, their spins, magnetic fields, etc.), and (ii) the occurrence rate of mergers in the Galaxy and in the local Universe. Masses of NSs in binaries are known with a rather good accuracy of 10% or better from, e.g., pulsar studies [400]; see also [211] for a recent update of NS mass measurements.

The case is not so good with the rate of coalescence of relativistic binary stars. Unfortunately, there is no way to derive it from first principles – neither the formation rate of the progenitor binaries for compact double stars nor stellar evolution are known well enough. However, the situation is not completely hopeless, especially in the case of double NS systems. Natural appearance of rotating NSs with magnetic fields as radio pulsars allows searching for binary pulsars with secondary compact companion using powerful methods of modern radio astronomy (for example, in dedicated pulsar surveys such as the Parkes multi-beam pulsar survey [244, 98]).

Based on the observational statistics of the Galactic binary pulsars with another NS companion, one can evaluate the Galactic rate of binary NS formation and merging [313, 270, 189]. On the other hand, a direct simulation of binary star evolution in the Galaxy (the *population synthesis* method) can also predict the formation and merger rates of close compact binaries as a function of (numerous) parameters of binary star formation and evolution. It is important and encouraging that both estimates (observational, as inferred from recent measurements of binary pulsars [48, 180], and theoretical from the population synthesis; see Section 6) now give very close estimates for the double NS star merger rate in the Galaxy of about one event per 10,000 years. No binary BH or NS + BH systems have been found so far, so merger rates of compact binaries with BHs have been evaluated as yet only from population synthesis studies.

Binaries with WDs

The interest in these binaries stems from several circumstances. First, they are considered as testbeds for gravitational wave physics. Second, with them SNe Ia are associated. SNe Ia are being used as the primary standard candle sources for the determination of the cosmological parameters Ω and Λ (see, e.g., [347, 307]). A comparison of SN Ia rates (for the different models of their progenitors) with observations may, in principle, shed light on both the star formation history and on the nature of the progenitors (see, e.g., [470, 243, 107]). The counts of distant SNe could be used to constrain cosmological parameters (see, e.g., [359]). Finally, close binaries with WDs are among the most promising verification binaries for LISA [390].

In this paper we shall concentrate on the formation and evolution of binary compact stars most relevant for GW studies. The paper is organized as follows. We start in Section 2 with a review of the main observational data on double NSs, especially measurements of masses of NSs and BHs, which are most important for the estimate of the amplitude of the expected GW signal. We briefly discuss the empirical methods to determine double NS coalescence rate. The basic principles of binary stellar evolution are discussed in Section 3. Then, in Section 4 we describe the evolution of massive binary stars. We then discuss the Galactic rate of formation of binaries with NSs and BHs in Section 5. Theoretical estimates of detection rates for mergers of binary relativistic stars are discussed in Section 6. Further we proceed to the analysis of formation of short-period binaries with WD components in Section 7, and consider observational data on binary white dwarfs in Section 8. A model for the evolution of interacting double-degenerate systems is presented in Section 9. In Section 10 we describe gravitational waves from compact binaries with white-dwarf components. Sections 11 and 12 are devoted, respectively, to the model of optical and X-ray emission of AM CVn-stars and to their subsample potentially observed both in electromagnetic and gravitational waves. Our conclusions follow in Section 13.

2 Observations of Double Neutron Stars

2.1 Compact binaries with neutron stars

Double NSs have been discovered because one of the components of the binary is observed as a radio pulsar. The precise pulsar timing allows one to search for a periodic variation due to the binary motion. This technique is reviewed in detail by Lorimer [236]; applications of pulsar timing for general relativity tests are reviewed by Stairs [386].

Basically, pulsar timing provides the following Keplerian orbital parameters of the binary system: the binary orbital period P_b as measured from periodic Doppler variations of the pulsar spin, the projected semimajor axis $x = a \sin i$ as measured from the semi-amplitude of the pulsar radial velocity curve (i is the binary inclination angle defined such that $i = 0$ for face-on systems), the orbital eccentricity e as measured from the shape of the pulsar radial velocity curve, and the longitude of periastron ω at a particular epoch T_0 . The first two parameters allow one to construct the *mass function* of the secondary companion,

$$f(M_p, M_c) = \frac{4\pi^2 x^3}{P_b^2 T_\odot} = \frac{(M_c \sin i)^3}{(M_c + M_p)^2}. \quad (1)$$

In this expression, x is measured in light-seconds, $T_\odot \equiv GM_\odot/c^3 = 4.925490947 \mu\text{s}$, and M_p and M_c denote masses of the pulsar and its companion, respectively. This function gives the strict lower limit on the mass of the unseen companion. However, assuming the pulsar mass to have the typical value of a NS mass (for example, confined between the lowest measured NS mass $1.25 M_\odot$ for PSR J0737-303B [239] and the maximum measured NS mass of $2.1 M_\odot$ in the NS–WD binary PSR J0751+1807 [290]), one can estimate the mass of the secondary star even without knowing the binary inclination angle i .

Long-term pulsar timing allows measurements of several relativistic phenomena: the advance of periastron $\dot{\omega}$, the redshift parameter γ , the Shapiro delay within the binary system qualified through post-Keplerian parameters r , s , and the binary orbit decay \dot{P}_b . From the post-Keplerian parameters the individual masses M_p , M_c and the binary inclination angle i can be calculated [45].

Of the post-Keplerian parameters of binary pulsars, the periastron advance rate is usually measured most readily. Assuming it to be entirely due to general relativity, the total mass of the system can be evaluated:

$$\dot{\omega} = 3 \left(\frac{2\pi}{P_b} \right)^{5/3} \frac{T_\odot^{2/3} (M_c + M_p)^{2/3}}{(1 - e^2)}. \quad (2)$$

High values of the derived total mass of the system ($\gtrsim 2.5 M_\odot$) suggests the presence of another NS or even BH².

If the individual masses, binary period, and eccentricity of a compact binary system are known, it is easy to calculate the time it takes for the binary companions to coalesce due to GW emission using the quadrupole formula for GW emission [308] (see Section 3.1.4 for more detail):

$$\tau_{\text{GW}} \approx 4.8 \times 10^{10} \text{ yr} \left(\frac{P_b}{\text{d}} \right)^{8/3} \left(\frac{\mu}{M_\odot} \right)^{-1} \left(\frac{M_c + M_p}{M_\odot} \right)^{-2/3} (1 - e^2)^{7/2}. \quad (3)$$

Here $\mu = M_p M_c / (M_p + M_c)$ the reduced mass of the binary. Some observed and derived parameters of known compact binaries with NSs are collected in Tables 2 and 3.

²Unless the companion is directly observable and its mass can be estimated by other means.

PSR	P [ms]	P_b [d]	$a_1 \sin i$ [lt-s]	e	$\dot{\omega}$ [deg yr ⁻¹]	\dot{P}_b [$\times 10^{-12}$]	Ref.
J0737-3039A	22.70	0.102	1.42	0.088	16.88	-1.24	[48]
J0737-3039B	2773	—	—	—	—	—	[239]
J1518+4904	40.93	8.634	20.04	0.249	0.011	?	[289]
B1534+12	37.90	0.421	3.73	0.274	1.756	-0.138	[455, 388]
J1756-2251	28.46	0.320	2.75	0.181	2.585	?	[97]
J1811-1736	104.18	18.779	34.78	0.828	0.009	< 30	[240]
J1906+074	144.07	0.116	1.42	0.085	7.57	?	[237]
B1913+16	59.03	0.323	2.34	0.617	4.227	-2.428	[153]
B2127+11C	30.53	0.335	2.52	0.681	4.457	-3.937	[7, 332]

Table 2: *Observed parameters of double neutron star binaries.*

PSR	$f(m)$ [M_\odot]	$M_c + M_p$ [M_\odot]	$\tau_c = P/(2\dot{P})$ [Myr]	τ_{GW} [Myr]
J0737-3039A	0.29	2.58	210	87
J0737-3039B	—	—	50	—
J1518+4904	0.12	2.62		9.6×10^6
B1534+12	0.31	2.75	248	2690
J1756-2251	0.22	2.57	444	1690
J1811-1736	0.13	2.6		1.7×10^6
J1906+074	0.11	2.61	0.112	300
B1913+16	0.13	2.83	108	310
B2127+11C	0.15	2.71	969	220

Table 3: *Derived parameters of double neutron star binaries*

2.2 How frequent are double NS coalescences?

As it is seen from Table 3, only six double NS systems presently known will coalesce over a time interval shorter than ≈ 10 Gyr: J0737–3039A, B1534+12, J1756–2251, J1906+074, B1913+16, and B2127+11C. Of these six systems, one (PSR B2127+11C) is located in the globular cluster M15. This system may have a different formation history, so usually it is not included in the analysis of the coalescence rate of Galactic double compact binaries. The formation and evolution of relativistic binaries in dense stellar systems is reviewed elsewhere [26]. For a recent general review of pulsars in globular clusters see also [51].

The ordinary way of estimating the double NS merger rate from binary pulsar statistics is based on the following extrapolation [270, 313]. Suppose we observe i classes of Galactic binary pulsars. Taking into account various selection effects of pulsar surveys (see, e.g., [269, 189]), the Galactic number of pulsars N_i in each class can be evaluated. To compute the Galactic merger rate of double NS binaries, we need to know the time since the birth of the NS observed as a pulsar in the given binary system. This time is the sum of the observed characteristic pulsar age τ_c and the time required for the binary system to merge due to GW orbit decay τ_{GW} . With the exception of PSR J0737–3039B and the recently discovered PSR J1906+074, pulsars that we observe in binary NS systems are old recycled pulsars which were spun-up by accretion from the secondary companion to the period of several tens of ms (see Table 2). Thus their characteristic ages can be estimated as the time since termination of spin-up by accretion (for the younger pulsar PSR J0737–3039B this time can be also computed as the dynamical age of the pulsar, $P/(2\dot{P})$, which gives essentially the same result).

Then the merger rate \mathcal{R}_i can be calculated as $\mathcal{R}_i \sim N_i/(\tau_c + \tau_{\text{GW}})$ (summed over all binary pulsars). The detailed analysis [189] indicates that the Galactic merger rate of double NSs is mostly determined by pulsars with faint radio luminosity and short orbital periods. Presently, it is the nearby (600 pc) double-pulsar system PSR J0737–3039 with a short orbital period of 2.4 hr [48] that mostly determines the *empirical* estimate of the merger rate. According to Kim et al. [190], “the most likely values of DNS merger rate lie in the range 3–190 per Myr depending on different pulsar models”. The estimates by *population synthesis codes* are still plagued by uncertainties in statistics of binaries, in modeling binary evolution and supernovae. The most optimistic “theoretical” predictions amount to $\simeq 300$ Myr $^{-1}$ [420, 22].

Recently, the bursting radio source GCRT J1745–3009 in the direction to the Galactic centre was proposed to be a possible double NS binary [407]. The source was found to emit a series of radio bursts with high brightness temperature, of typical duration ~ 10 min, with an apparent periodic pattern of ~ 77 min [156]. Confirmation of the binary NS nature of transient radio sources like GCRT J1745–3009 would be important to get a more precise estimate of the Galactic coalescence rate of double NS.

The extrapolation beyond the Galaxy is usually done by scaling the Galactic merger rate to the volume from which the merger events can be detected with given GW detector’s sensitivity. The scaling factor widely used is the ratio between the B-band luminosity density in the local Universe, correlating with the star-formation rate, and the B-band luminosity of the Galaxy [313, 182]. One can also use for this purpose the direct ratio of the Galactic star formation rate $\text{SFR}_{\text{G}} \simeq 3 M_{\odot} \text{ yr}^{-1}$ [255, 401] to the star formation rate on the local Universe $\text{SFR}_{\text{loc}} \simeq 0.03 M_{\odot} \text{ yr}^{-1}$ [306, 371]. These estimates yield the relation

$$\mathcal{R}_{\text{V}} \approx 0.01 \mathcal{R}_{\text{G}} [\text{Mpc}^{-3}]. \quad (4)$$

So for the Galactic merger rate $\mathcal{R}_{\text{G}} \sim 10^{-4} \text{ yr}^{-1}$ a very optimistic detection rate for binary NSs of about once per 1–2 years of observations by the first-generation GW detectors is predicted [48]. This estimate is still uncertain, mostly due to poor knowledge of the luminosity function for faint radio pulsars [190].

Recently, the first results of the search for GWs from coalescing binary systems in the Milky Way and the Magellanic Clouds using data taken by two of the three LIGO interferometers [3] established an observational upper limit to the Galactic binary NS coalescence rate of $\mathcal{R}_G < 1.7 \times 10^2 \text{ yr}^{-1}$. With increasing sensitivity of GW detectors, this limit will be much improved in the nearest future.

2.3 On the connection between GRBs and compact binary mergers

Here we wish to mention the possibility of observational manifestations of NS/BH binary mergers other than the violent GW emission. First of all, it is the connection of relativistic binary mergers to some subclasses of cosmic GRBs. That catastrophic events like the coalescence of binary NSs or BHs can be related to GRBs has been suggested for quite a long time. First ideas can be found in the papers of Sergei Blinnikov and his coauthors [39] and Bohdan Paczyński [302] with subsequent studies in [90, 303, 258, 188], etc.

Now these ideas gained strong observational support from the accurate localization of short GRBs with hard spectrum by the Swift and HETE-II space missions [116, 40, 31, 109]. The short-hard subclass of GRBs includes up to 30% of all GRBs [299]. The most important recent discovery is that these GRBs occur both in late-type [66, 146] and early-type galaxies [31, 18], suggesting old stellar population progenitors. This is in sharp contrast to long GRBs, some of which are definitely associated with peculiar type Ib/c supernovae produced by the core collapse of massive stars [145, 389].

The principal observational facts about several well-localized short GRBs (see [29] for more discussion) are:

1. They occur at cosmological redshifts from 0.160 to 1.8 and may constitute up to 20% of the local short GRB population detected by the BATSE³ experiment on board of the Compton Gamma-ray Observatory.
2. The isotropic energy release is typically lower than in long GRBs (from $\sim 10^{48}$ erg to 4×10^{51} erg).
3. The opening angle of the ejecta in these GRBs is on average larger than in long GRBs [50].
4. Short GRBs are found both in elliptical and star forming galaxies. Statistical analysis suggests that the occurrence rate of short bursts is roughly equal in early-type and late-type galaxies [29].

The short GRB rate inferred from these observations [317], $\mathcal{R}_{\text{sGRB}} \sim (10-30) \frac{3}{70} \text{ Gpc}^{-3} \text{ yr}^{-1}$ agrees with the double NS merger rate derived from binary pulsar statistics. Depending on the unknown beaming factor a possible upper limit of about $10^5 \text{ events yr}^{-1} \text{ Gpc}^{-3}$ was obtained in [264]. That paper also extensively discusses the application of the rate of short GRBs to LIGO/VIRGO detections of double NS binary mergings if they are associated with short GRBs, and gives very good prospects for the Advanced LIGO sensitivity (up to hundreds detections per year). However, recent deep optical observations of several short GRBs provide evidence for their association with very faint galaxies [30], suggesting the intrinsic luminosity of a significant part of short GRBs to be much higher than $10^{48} - 10^{49}$ erg, as inferred from observations of close short GRBs by [264], close to that of classical long GRBs $\sim 10^{51} - 10^{52}$ erg. Taking this finding into account decreases the expected detection rate of NS mergers (if they are associated with short GRBs) down to several events per year by the Advanced LIGO detector [30].

We also emphasize the agreement of the observational estimates with population synthesis calculations of binary mergers in galaxies of different types [14, 305, 21, 46, 24]. The analysis of

³Burst And Transient Source Experiment [271].

luminosity function and statistics of short GRBs from the BATSE catalog [317] implies a delay relative to the star formation history, which can favour double NS systems dynamically formed in stellar clusters as progenitors [150, 121]. Theoretical issues related to the generation of short hard GRBs from binary NS and NS–BH mergers are discussed in [212, 213, 295].

Now let us see what theory says about the formation, evolution, and detection rates of close compact binaries and their properties.

3 Basic Principles of Binary Star Evolution

Beautiful general reviews of the topic can be found in [33, 425]. Here we restrict ourselves to recalling several facts about binary evolution which are most relevant to the formation and evolution of compact binaries. The readers who have experience in the field can skip this section.

3.1 Keplerian binary system and radiation back reaction

We start with some basic facts about Keplerian motion in a binary system and the simplest case of evolution of two point masses due to gravitational radiation losses. The stars are highly condensed objects, so their treatment as point masses is usually adequate for the description of their interaction in the binary. Furthermore, Newtonian gravitation theory is sufficient for this purpose as long as the orbital velocities are small in comparison with the speed of light c . The systematic change of the orbit caused by the emission of gravitational waves will be considered in a separate paragraph below.

3.1.1 Keplerian motion

Let us consider two point masses M_1 and M_2 orbiting each other under the force of gravity. It is well known (see [208]) that this problem is equivalent to the problem of a single body with mass μ moving in an external gravitational potential. The value of the external potential is determined by the total mass of the system

$$M = M_1 + M_2. \quad (5)$$

The reduced mass μ is

$$\mu = \frac{M_1 M_2}{M}. \quad (6)$$

The body μ moves in an elliptic orbit with eccentricity e and major semi-axis a . The orbital period P and orbital frequency $\Omega = 2\pi/P$ are related to M and a by Kepler's third law

$$\Omega^2 = \left(\frac{2\pi}{P}\right)^2 = \frac{GM}{a^3}. \quad (7)$$

This relationship is valid for any eccentricity e .

Individual bodies M_1 and M_2 move around the barycentre of the system in elliptic orbits with the same eccentricity e . The major semi-axes a_i of the two ellipses are inversely proportional to the masses

$$\frac{a_1}{a_2} = \frac{M_2}{M_1}, \quad (8)$$

and satisfy the relationship $a = a_1 + a_2$. The position vectors of the bodies from the system's barycentre are $\vec{r}_1 = M_2 \vec{r}/(M_1 + M_2)$ and $\vec{r}_2 = -M_1 \vec{r}/(M_1 + M_2)$, where $\vec{r} = \vec{r}_1 - \vec{r}_2$ is the relative position vector. Therefore, the velocities of the bodies with respect to the system's barycentre are related by

$$-\frac{\vec{V}_1}{\vec{V}_2} = \frac{M_2}{M_1}, \quad (9)$$

and the relative velocity is $\vec{V} = \vec{V}_1 - \vec{V}_2$.

The total conserved energy of the binary system is

$$E = \frac{M_1 \vec{V}_1^2}{2} + \frac{M_2 \vec{V}_2^2}{2} - \frac{GM_1 M_2}{r} = \frac{\mu \vec{V}^2}{2} - \frac{GM_1 M_2}{r} = -\frac{GM_1 M_2}{2a}, \quad (10)$$

where r is the distance between the bodies. The orbital angular momentum vector is perpendicular to the orbital plane and can be written as

$$\vec{J}_{\text{orb}} = M_1 \vec{V}_1 \times \vec{r}_1 + M_2 \vec{V}_2 \times \vec{r}_2 = \mu \vec{V} \times \vec{r}. \quad (11)$$

The absolute value of the orbital angular momentum is

$$|\vec{J}_{\text{orb}}| = \mu \sqrt{GMa(1 - e^2)}. \quad (12)$$

For circular binaries with $e = 0$ the distance between orbiting bodies does not depend on time,

$$r(t, e = 0) = a,$$

and is usually referred to as orbital separation. In this case, the velocities of the bodies, as well as their relative velocity, are also time-independent,

$$V \equiv |\vec{V}| = \Omega a = \sqrt{GM/a}, \quad (13)$$

and the orbital angular momentum becomes

$$|\vec{J}_{\text{orb}}| = \mu V a = \mu \Omega a^2. \quad (14)$$

3.1.2 Gravitational radiation from a binary

The plane of the orbit is determined by the orbital angular momentum vector \vec{J}_{orb} . The line of sight is defined by a unit vector \vec{n} . The binary inclination angle i is defined by the relation $\cos i = (\vec{n}, \vec{J}_{\text{orb}}/J_{\text{orb}})$ such that $i = 90^\circ$ corresponds to a system visible edge-on.

Let us start from two point masses M_1 and M_2 in a circular orbit. In the quadrupole approximation [209], the two polarization amplitudes of GWs at a distance r from the source are given by

$$h_+ = \frac{G^{5/3}}{c^4} \frac{1}{r} 2(1 + \cos^2 i)(\pi f M)^{2/3} \mu \cos(2\pi f t), \quad (15)$$

$$h_\times = \pm \frac{G^{5/3}}{c^4} \frac{1}{r} 4 \cos i (\pi f M)^{2/3} \mu \sin(2\pi f t). \quad (16)$$

Here $f = \Omega/\pi$ is the frequency of the emitted GWs (twice the orbital frequency). Note that for a fixed distance r and a given frequency f , the GW amplitudes are fully determined by $\mu M^{2/3} = \mathcal{M}^{5/3}$, where the combination

$$\mathcal{M} \equiv \mu^{3/5} M^{2/5}$$

is called the ‘‘chirp mass’’ of the binary. After averaging over the orbital period (so that the squares of periodic functions are replaced by 1/2) and the orientations of the binary orbital plane, one arrives at the averaged (characteristic) GW amplitude

$$h(f, \mathcal{M}, r) = (\langle h_+^2 \rangle + \langle h_\times^2 \rangle)^{1/2} = \left(\frac{32}{5}\right)^{1/2} \frac{G^{5/3}}{c^4} \frac{\mathcal{M}^{5/3}}{r} (\pi f)^{2/3}. \quad (17)$$

3.1.3 Energy and angular momentum loss

In the approximation and under the choice of coordinates that we are working with, it is sufficient to use the Landau–Lifshitz gravitational pseudo-tensor [209] when calculating the gravitational waves energy and flux. (This calculation can be justified with the help of a fully satisfactory gravitational energy-momentum tensor that can be derived in the field theory formulation of general

relativity [11]). The energy dE carried by a gravitational wave along its direction of propagation per area dA per time dt is given by

$$\frac{dE}{dA dt} \equiv F = \frac{c^3}{16\pi G} \left[\left(\frac{\partial h_+}{\partial t} \right)^2 + \left(\frac{\partial h_\times}{\partial t} \right)^2 \right]. \quad (18)$$

The energy output dE/dt from a localized source in all directions is given by the integral

$$\frac{dE}{dt} = \int F(\theta, \phi) r^2 d\Omega. \quad (19)$$

Replacing

$$\left(\frac{\partial h_+}{\partial t} \right)^2 + \left(\frac{\partial h_\times}{\partial t} \right)^2 = 4\pi^2 f^2 h^2(\theta, \phi)$$

and introducing

$$h^2 = \frac{1}{4\pi} \int h^2(\theta, \phi) d\Omega,$$

we write Equation (19) in the form

$$\frac{dE}{dt} = \frac{c^3}{G} (\pi f)^2 h^2 r^2. \quad (20)$$

Specifically for a binary system in a circular orbit, one finds the energy loss from the system (sign minus) with the help of Equations (20) and (17):

$$\frac{dE}{dt} = - \left(\frac{32}{5} \right) \frac{G^{7/3}}{c^5} (\mathcal{M}\pi f)^{10/3}. \quad (21)$$

This expression is exactly the same one that can be obtained directly from the quadrupole formula [209],

$$\frac{dE}{dt} = - \frac{32}{5} \frac{G^4}{c^5} \frac{M_1^2 M_2^2 M}{a^5}, \quad (22)$$

rewritten using the definition of the chirp mass and Kepler's law. Since energy and angular momentum are continuously carried away by gravitational radiation, the two masses in orbit spiral towards each other, thus increasing their orbital frequency Ω . The GW frequency $f = \Omega/\pi$ and the GW amplitude h are also increasing functions of time. The rate of the frequency change is⁴

$$\dot{f} = \left(\frac{96}{5} \right) \frac{G^{5/3}}{c^5} \pi^{8/3} \mathcal{M}^{5/3} f^{11/3}. \quad (23)$$

In spectral representation, the flux of energy per unit area per unit frequency interval is given by the right-hand-side of the expression

$$\frac{dE}{dA df} = \frac{c^3}{G} \frac{\pi f^2}{2} \left(\left| \tilde{h}(f)_+ \right|^2 + \left| \tilde{h}(f)_\times \right|^2 \right) \equiv \frac{c^3}{G} \frac{\pi f^2}{2} S_h^2(f), \quad (24)$$

where we have introduced the spectral density $S_h^2(f)$ of the gravitational wave field h . In the case of a binary system, the quantity S_h is calculable from Equations (15, 16):

$$S_h = \frac{G^{5/3}}{c^3} \frac{\pi}{12} \frac{\mathcal{M}^{5/3}}{r^2} \frac{1}{(\pi f)^{7/3}}. \quad (25)$$

⁴A signal with such an increasing frequency is reminiscent of the chirp of a bird. This explains the origin of the term "chirp mass" for the parameter \mathcal{M} which fully determines the GW frequency and amplitude behaviour.

3.1.4 Binary coalescence time

A binary system in a circular orbit loses energy according to Equation (21). For orbits with non-zero eccentricity e , the right-hand-side of this formula should be multiplied by the factor

$$f(e) = \left(1 + \frac{73}{24}e^2 + \frac{37}{96}e^4\right)(1 - e^2)^{-7/2}$$

(see [308]). The initial binary separation a_0 decreases and, assuming Equation (22) is always valid, it should vanish in a time

$$t_0 = \frac{c^5}{G^3} \frac{5a_0^4}{256M^2\mu} = \frac{5c^5}{256} \frac{(P_0/2\pi)^{8/3}}{(GM)^{5/3}} \approx (9.8 \times 10^6 \text{ yr}) \left(\frac{P_0}{1 \text{ h}}\right)^{8/3} \left(\frac{\mathcal{M}}{M_\odot}\right)^{-5/3}. \quad (26)$$

As we noted above, gravitational radiation from the binary depends on the chirp mass \mathcal{M} , which can also be written as $\mathcal{M} \equiv M\eta^{3/5}$, where η is the dimensionless ratio $\eta = \mu/M$. Since $\eta \leq 1/4$, one has $\mathcal{M} \lesssim 0.435M$. For example, for two NS with equal masses $M_1 = M_2 = 1.4 M_\odot$, the chirp mass is $\mathcal{M} \approx 1.22 M_\odot$. This explains the choice of normalization in Equation (26).

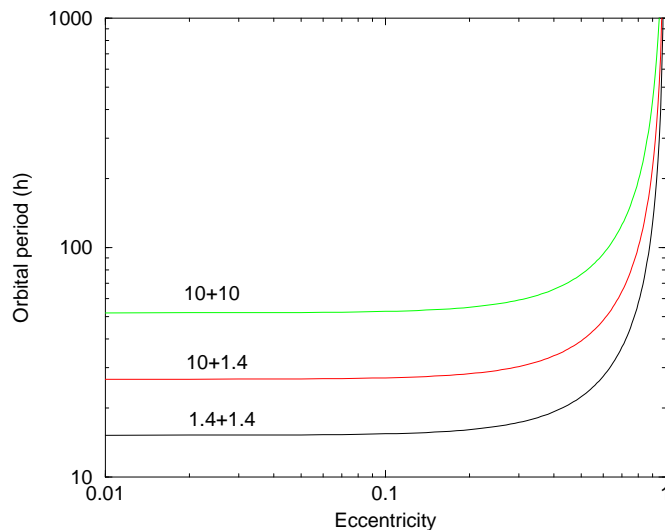


Figure 3: *The maximum initial orbital period (in hours) of two point masses which will coalesce due to gravitational wave emission in a time interval shorter than 10^{10} yr, as a function of the initial eccentricity e_0 . The lines are calculated for $10 M_\odot + 10 M_\odot$ (BH + BH), $10 M_\odot + 1.4 M_\odot$ (BH + NS), and $1.4 M_\odot + 1.4 M_\odot$ (NS + NS).*

The coalescence time for an initially eccentric orbit with $e_0 \neq 0$ and separation a_0 is shorter than the coalescence time for a circular orbit with the same initial separation a_0 [308]:

$$t_c(e_0) = t_0 f(e_0), \quad (27)$$

where the correction factor $f(e_0)$ is

$$f(e_0) = \frac{48}{19} \frac{(1 - e_0^2)^4}{e_0^{48/19} (1 + \frac{121}{304}e_0^2)^{3480/2299}} \int_0^{e_0} \frac{(1 + \frac{121}{304}e^2)^{1181/2299}}{(1 - e^2)^{3/2}} e^{29/19} de. \quad (28)$$

To merge in a time interval shorter than 10 Gyr the binary should have a small enough initial orbital period $P_0 \leq P_{cr}(e_0, \mathcal{M})$ and, accordingly, a small enough initial semimajor axis $a_0 \leq$

$a_{cr}(e_0, \mathcal{M})$. The critical orbital period is plotted as a function of the initial eccentricity e_0 in Figure 3. The lines are plotted for three typical sets of masses: two neutron stars with equal masses ($1.4 M_\odot + 1.4 M_\odot$), a black hole and a neutron star ($10 M_\odot + 1.4 M_\odot$), and two black holes with equal masses ($10 M_\odot + 10 M_\odot$). Note that in order to get a significantly shorter coalescence time, the initial binary eccentricity should be $e_0 \geq 0.6$.

3.1.5 Magnetic stellar wind

In the case of low-mass binary evolution, there is another important physical mechanism responsible for the removal of orbital angular momentum, in addition to GW emission discussed above. This is the magnetic stellar wind (MSW), or magnetic braking, which is thought to be effective for main-sequence G-M dwarfs with convective envelopes, i.e. in the mass interval $0.3 - 1.5 M_\odot$. The upper mass limit corresponds to the disappearance of a deep convective zone, while the lower mass limit stands for fully convective stars. In both cases a dynamo mechanism, responsible for enhanced magnetic activity, is thought to be ineffective. The idea behind angular momentum loss (AML) by magnetically coupled stellar wind is that the stellar wind is compelled by magnetic field to corotate with the star to rather large distances where it carries away large specific angular momentum [370]. Thus, it appears possible to take away substantial angular momentum without evolutionary significant mass-loss in the wind. The concept of an MSW was introduced into analyses of the evolution of compact binaries by Verbunt and Zwaan [435] when it became evident that momentum loss by GWs alone is unable to explain the observed mass-transfer rates in cataclysmic variables. The latter authors based their reasoning on observations of the spin-down of single G-dwarfs in stellar clusters with age [378] $V \propto t^{-1/2}$ (the Skumanich law). Applying this to a binary component and assuming tidal locking between the star's axial rotation and orbital revolution, one arrives at the rate of angular momentum loss via an MSW,

$$\frac{\dot{J}_{\text{MSW}}}{J_{\text{orb}}} \sim -\frac{R_o^4}{M_c} \frac{GM^2}{a^5}, \quad (29)$$

where R_o is the radius of the optical companion and M_c is the mass of the compact star.

Radii of stars filling their Roche lobes should be proportional to binary separations, $R_o \propto a$, which means that the characteristic time of orbital angular momentum removal by an MSW is $\tau_{\text{MSW}} \equiv (\dot{J}_{\text{MSW}}/J_{\text{orb}})^{-1} \propto a$. This should be compared with AML by GWs with $\tau_{\text{GW}} \propto a^4$. Clearly, the MSW (if it operates) is more effective in removing angular momentum from binary system at larger separations (orbital periods), and at small orbital periods GWs always dominate. Magnetic braking is especially important in CVs and in LMXBs with orbital periods exceeding several hours and is the driving mechanism for mass accretion onto the compact component.

We should note that the above prescription for an MSW is still debatable, since it is based on extrapolation of stellar rotation rates over several orders of magnitude – from slowly rotating field stars to rapidly spinning components of close binaries. There are strong indications that actual magnetic braking for compact binaries may be much weaker than predictions based on the Skumanich law (see, e.g., [323] for recent discussion and references).

3.2 Mass transfer modes and mass loss in binary systems

GW emission is the sole factor responsible for the change of orbital parameters of a pair of compact (degenerate) stars. However, at the early stages of binary evolution, it is the mass transfer between the components and the loss of matter and its orbital momentum that play a dominant dynamical role. Strictly speaking, these processes should be treated hydro-dynamically and they require complicated numerical calculations. However, binary evolution can also be described semiquantitatively, using a simplified description in terms of point-like bodies. The change of their integrated physical

quantities, such as masses, orbital angular momentum, etc. governs the evolution of the orbit. This description turns out to be successful in reproducing the results of more rigorous numerical calculations (see, e.g., [294] for more detail and references). In this approach, the key role is allocated to the total orbital angular momentum J_{orb} of the binary.

Let star 2 lose matter at a rate $\dot{M}_2 < 0$ and let β ($0 \leq \beta \leq 1$) be the fraction of the ejected matter which leaves the system (the rest falls on the first star), i.e. $\dot{M}_1 = -(1-\beta)\dot{M}_2 \geq 0$. Consider circular orbits with orbital angular momentum given by Equation (14). Differentiate both parts of Equation (14) by time t and exclude $d\Omega/dt$ with the help of Kepler's third law (7). This gives us the rate of change of the orbital separation:

$$\frac{\dot{a}}{a} = -2 \left(1 + (\beta - 1) \frac{M_2}{M_1} - \frac{\beta}{2} \frac{M_2}{M} \right) \frac{\dot{M}_2}{M_2} + 2 \frac{\dot{J}_{\text{orb}}}{J_{\text{orb}}}. \quad (30)$$

In the previous equation \dot{a} and \dot{M} are not independent variables if the donor fills its Roche lobe. One defines the mass transfer as conservative if both $\beta = 0$ and $\dot{J}_{\text{orb}} = 0$. The mass transfer is called non-conservative if at least one of these conditions is violated.

It is important to distinguish some specific cases (modes) of mass transfer:

1. conservative mass transfer,
2. non-conservative Jeans mode of mass loss (or fast wind mode),
3. non-conservative isotropic re-emission,
4. sudden mass loss from one of the component during supernova explosion, and
5. common-envelope stage.

As specific cases of angular momentum loss we consider GW emission (see Section 3.1.3 and 3.1.4) and a magnetically coupled stellar wind (see Section 3.1.5) which drive the orbital evolution for short-period binaries. For non-conservative modes, one can also introduce some sub-cases, such as, for example, a ring-like mode in which a circumbinary ring of expelled matter is being formed (see, e.g., [380]). Here, we will not go into the details of such sub-cases.

3.2.1 Conservative accretion

In the case of conservative accretion, matter from M_2 is fully deposited onto M_1 . The transfer process preserves the total mass ($\beta = 0$) and the orbital angular momentum of the system. It follows from Equation (30) that

$$M_1 M_2 \sqrt{a} = \text{const},$$

so that the initial and final binary separations are related as

$$\frac{a_f}{a_i} = \left(\frac{M_{1i} M_{2i}}{M_{1f} M_{2f}} \right)^2. \quad (31)$$

The well-known ‘‘rule of thumb’’ for this case says that the orbit shrinks when the more massive component loses matter, and the orbit widens in the opposite situation. During such a mass exchange, the orbital separation passes through a minimum, if the masses become equal in course of mass transfer.

3.2.2 The Jeans (fast wind) mode

In this mode the ejected matter completely escapes from the system, that is, $\beta = 1$. The escape of matter can take place either in a spherically symmetric way or in the form of bipolar jets moving from the system at high velocity. In both cases, matter carries away some amount of the total orbital momentum proportional to the orbital angular momentum of the mass losing star $J_2 = (M_1/M)J_{\text{orb}}$ (we neglect a possible proper rotation of the star, see [425]). For the loss of orbital momentum \dot{J}_{orb} it is reasonable to take

$$\dot{J}_{\text{orb}} = \frac{\dot{M}_2}{M_2} J_2. \quad (32)$$

In the case $\beta = 1$, Equation (30) can be written as

$$\frac{(\dot{\Omega}a^2)}{\Omega a^2} = \frac{\dot{J}_{\text{orb}}}{J_{\text{orb}}} - \frac{M_1 \dot{M}_2}{MM_2}. \quad (33)$$

Then Equation (33) in conjunction with Equation (32) gives $\Omega a^2 = \text{const}$, that is, $\sqrt{GaM} = \text{const}$. Thus, as a result of such a mass loss, the change in orbital separation is

$$\frac{a_f}{a_i} = \frac{M_i}{M_f}. \quad (34)$$

Since the total mass decreases, the orbit always widens.

3.2.3 Isotropic re-emission

The matter lost by star 2 can first accrete to star 1, and then, a fraction β of the accreted matter, can be expelled from the system. This happens, for instance, when a massive star transfers matter to a compact star on the thermal timescale ($< 10^6$ years). The accretion luminosity may exceed the Eddington luminosity limit, and the radiation pressure pushes the infalling matter away from the system, in a manner similar to the spectacular example of the SS 433 binary system. Another examples may be systems with helium stars transferring mass onto relativistic objects [234, 117] or hot white dwarf components of cataclysmic variables losing mass by optically thick winds [187]. The same algorithm may be applied to the time-averaged mass loss from novae [466]. In this mode of mass-transfer, the binary orbital momentum carried away by the expelled matter is determined by the orbital momentum of the accreting star M_1 , rather than by the orbital momentum of the mass-losing star M_2 . The orbital momentum loss can be written as

$$\dot{J}_{\text{orb}} = \beta \frac{\dot{M}_2}{M_1} J_1, \quad (35)$$

where $J_1 = (M_2/M)J_{\text{orb}}$ is the orbital momentum of the star M_1 . In the limiting case when all the mass attracted by M_1 is expelled from the system, $\beta = 1$, Equation (35) simplifies to

$$\frac{\dot{J}_{\text{orb}}}{J_{\text{orb}}} = \frac{\dot{M}_2 M_2}{M_1 M}. \quad (36)$$

After substitution of this formula into Equation (30) and integration over time, one arrives at

$$\frac{a_f}{a_i} = \frac{M_i}{M_f} \left(\frac{M_{2i}}{M_{2f}} \right)^2 \exp \left(-2 \frac{M_{2i} - M_{2f}}{M_1} \right). \quad (37)$$

The exponential term makes this mode of the mass transfer very sensitive to the components mass ratio. If $M_1/M_2 \ll 1$, the separation a between the stars may decrease so much that the approximation of point masses becomes invalid. The tidal orbital instability (Darwin instability) may set in, and the compact star may start spiraling toward the companion star center (the common envelope stage; see Section 3.5 below).

3.3 Supernova explosion

A supernova explosion in a massive binary system occurs on a timescale much shorter than the orbital period, so the loss of mass is practically instantaneous. This case can be treated analytically (see, e.g., [38, 106, 257, 392, 140, 460, 42, 178, 397]).

In general, the loss of matter and radiation is non-spherical, so that the remnant of the supernova explosion (neutron star or black hole) acquires some recoil velocity called kick velocity \vec{w} . In a binary, the kick velocity should be added to the orbital velocity of the pre-supernova star.

The usual treatment proceeds as follows. Let us consider a pre-SN binary with initial masses M_1 and M_2 . The stars move in a circular orbit with orbital separation a_i and relative velocity \vec{V}_i . The star M_1 explodes leaving a compact remnant of mass M_c . The total mass of the binary decreases by the amount $\Delta M = M_1 - M_c$. It is usually assumed that the compact star acquires some additional velocity (kick velocity) \vec{w} (see detailed discussion in Section 3.4). Unless the binary is disrupted, it will end up in a new orbit with eccentricity e , major semi axis a_f , and angle θ between the orbital planes before and after the explosion. In general, the new barycentre will also receive some velocity, but we neglect this motion. The goal is to evaluate the parameters a_f , e , and θ .

It is convenient to work in an instantaneous reference frame centered on M_2 right at the time of explosion. The x -axis is the line from M_2 to M_1 , the y -axis points in the direction of \vec{V}_i , and the z -axis is perpendicular to the orbital plane. In this frame, the pre-SN relative velocity is $\vec{V}_i = (0, V_i, 0)$, where $V_i = \sqrt{G(M_1 + M_2)/a_i}$ (see Equation (13)). The initial total orbital momentum is $\vec{J}_i = \mu_i a_i (0, 0, -V_i)$. The explosion is considered to be instantaneous. Right after the explosion, the position vector of the exploded star M_1 has not changed: $\vec{r} = (a_i, 0, 0)$. However, other quantities have changed: $\vec{V}_f = (w_x, V_i + w_y, w_z)$ and $\vec{J}_f = \mu_f a_i (0, w_z, -(V_i + w_y))$, where $\vec{w} = (w_x, w_y, w_z)$ is the kick velocity and $\mu_f = M_c M_2 / (M_c + M_2)$ is the reduced mass of the system after explosion. The parameters a_f and e are being found from equating the total energy and the absolute value of orbital momentum of the initial circular orbit to those of the resulting elliptical orbit (see Equations (10, 14, 12)):

$$\mu_f \frac{V_f^2}{2} - \frac{GM_c M_2}{a_i} = -\frac{GM_c M_2}{2a_f}, \quad (38)$$

$$\mu_f a_i \sqrt{w_z^2 + (V_i + w_y)^2} = \mu_f \sqrt{G(M_c + M_2) a_f (1 - e^2)}. \quad (39)$$

For the resulting a_f and e one finds

$$\frac{a_f}{a_i} = \left[2 - \chi \left(\frac{w_x^2 + w_z^2 + (V_i + w_y)^2}{V_i^2} \right) \right]^{-1} \quad (40)$$

and

$$1 - e^2 = \chi \frac{a_i}{a_f} \left(\frac{w_z^2 + (V_i + w_y)^2}{V_i^2} \right), \quad (41)$$

where $\chi \equiv (M_1 + M_2)/(M_c + M_2) \geq 1$. The angle θ is defined by

$$\cos \theta = \frac{\vec{J}_f \cdot \vec{J}_i}{|\vec{J}_f| |\vec{J}_i|},$$

which results in

$$\cos \theta = \frac{V_i + w_y}{\sqrt{w_z^2 + (V_i + w_y)^2}}. \quad (42)$$

The condition for disruption of the binary system depends on the absolute value V_f of the final velocity, and on the parameter χ . The binary disrupts if its total energy defined by the

left-hand-side of Equation (38) becomes non-negative or, equivalently, if its eccentricity defined by Equation (41) becomes $e \geq 1$. From either of these requirements one derives the condition for disruption:

$$\frac{V_f}{V_i} \geq \sqrt{\frac{2}{\chi}}. \quad (43)$$

The system remains bound if the opposite inequality is satisfied. Equation (43) can also be written in terms of the escape (parabolic) velocity V_e defined by the requirement

$$\mu_f \frac{V_e^2}{2} - \frac{GM_c M_2}{a_i} = 0.$$

Since $\chi = M/(M - \Delta M)$ and $V_e^2 = 2G(M - \Delta M)/a_i = 2V_i^2/\chi$, one can write Equation (43) in the form

$$V_f \geq V_e. \quad (44)$$

The condition of disruption simplifies in the case of a spherically symmetric SN explosion, that is, when there is no kick velocity, $\vec{w} = 0$, and, therefore, $V_f = V_i$. In this case, Equation (43) reads $\chi \geq 2$, which is equivalent to $\Delta M \geq M/2$. Thus, the system unbinds if more than a half of the mass of the binary is lost. In other words, the resulting eccentricity

$$e = \frac{M_1 - M_c}{M_c + M_2} \quad (45)$$

following from Equations (40, 41), and $\vec{w} = 0$ becomes larger than 1, if $\Delta M > M/2$.

So far, we have considered an originally circular orbit. If the pre-SN star moves in an originally eccentric orbit, the condition of disruption of the system under symmetric explosion reads

$$\Delta M = M_1 - M_c > \frac{1}{2} \frac{r}{a_i},$$

where r is the distance between the components at the moment of explosion.

3.4 Kick velocity of neutron stars

The kick velocity imparted to a NS at birth is one of the major problems in the theory of stellar evolution. By itself it is an additional parameter, the introduction of which has been motivated first of all by high space velocities of radio pulsars inferred from the measurements of their proper motions and distances. Pulsars were recognized as a high-velocity Galactic population soon after their discovery in 1968 [125]. Shklovskii [376] put forward the idea that high pulsar velocities may result from asymmetric supernova explosions. Since then this hypothesis has been tested by pulsar observations, but no definite conclusions on its magnitude and direction have been obtained as yet.

Indeed, the distance to a pulsar is usually derived from the dispersion measure evaluation and crucially depends on the assumed model of the electron density distribution in the Galaxy. In the middle of the 1990s, Lyne and Lorimer [241] derived a very high mean space velocity of pulsars with known proper motion of about 450 km s^{-1} . This value was difficult to adopt without invoking an additional natal kick velocity of NSs. It was suggested [225] that the observed 2D pulsar velocity distribution found by Lyne and Lorimer [241] is reproduced if the absolute value of the (assumed to be isotropic) NS kick has a power-law shape,

$$f_{\text{LL}}(|\vec{w}|) \propto \frac{(|\vec{w}|/w_0)^{0.19}}{(1 + (|\vec{w}|/w_0)^{6.72})^{0.5}}, \quad (46)$$

with $w_0 \approx 400 \text{ km s}^{-1}$. However, using this formula or a Maxwellian distribution for NS kicks in population synthesis calculations [225] give similar results, and later we shall not distinguish these

kicks. The high mean space velocity of pulsars, consistent with earlier results by Lyne and Lorimer, was confirmed by the analysis of a larger sample of pulsars [147]. The recovered distribution of 3D velocities is well fit by a Maxwellian distribution with the mean value $w_0 = 400 \pm 40 \text{ km s}^{-1}$ and a 1D rms $\sigma = 265 \text{ km s}^{-1}$.

Possible physical reasons for natal NS kicks due to hydrodynamic effects in core-collapse supernovae are summarized in [207, 206]. Neutrino effects in the strong magnetic field of a young NS may be also essential in explaining kicks up to $\sim 100 \text{ km s}^{-1}$ [56, 83, 205]. Astrophysical arguments favouring a kick velocity are also summarized in [398]. To get around the theoretical difficulty of insufficient rotation of pre-supernova cores in single stars to produce rapidly spinning young pulsars, Spruit and Phinney [385] proposed that random off-center kicks can lead to a net spin-up of proto-NSs. In this model, correlations between pulsar space velocity and rotation are possible and can be tested in further observations.

Here we should note that the existence of some kick follows not only from the measurements of radio pulsar space velocities, but also from the analysis of binary systems with NSs. The impact of a kick velocity $\sim 100 \text{ km s}^{-1}$ explains the precessing binary pulsar orbit in PSR J0045–7319 [186]. The evidence of the kick velocity is seen in the inclined, with respect to the orbital plane, circumstellar disk around the Be star SS 2883 – an optical component of binary PSR B1259–63 [334].

Long-term pulse profile changes interpreted as geodetic precession are observed in the relativistic binary pulsars PSR 1913+16 [447], PSR B1534+12 [387], PSR J1141–6545 [151], and PSR J0737–3039B [49]. These observations indicate that in order to produce the misalignment between the orbital angular momentum and the neutron star spin, a component of the kick velocity perpendicular to the orbital plane is required [450, 453, 454]. This idea seems to gain observational support from recent thorough polarization measurements [173] suggesting alignment of the rotational axes with pulsar’s space velocity. Such an alignment acquired at birth may indicate the kick velocity directed preferably along the rotation of the proto-NS. For the first SN explosion in a close binary system this would imply the kick to be perpendicular to the orbital plane.

It is worth noticing that the analysis of the formation of the double relativistic pulsar PSR J0737–3039 [322] may suggest, from the observed low eccentricity of the system $e \simeq 0.09$, that a small (if any) kick velocity may be acquired if the formation of the second NS in the system is associated with the collapse of an ONeMg WD due to electron-captures. The symmetric nature of electron-capture supernovae was discussed in [319] and seems to be an interesting issue requiring further studies (see, e.g., [310, 204] for the analysis of the formation of NSs in globular clusters in the frame of this hypothesis). Note that electron-capture SNe are expected to be weak events, irrespective of whether they are associated with the core-collapse of a star which retained some original envelope or with the AIC of a WD [349, 192, 77].

We also note the hypothesis of Pfahl et al. [312], based on observations of high-mass X-ray binaries with long orbital periods ($\gtrsim 30 \text{ d}$) and low eccentricities ($e < 0.2$), that rapidly rotating precollapse cores may produce neutron stars with relatively small kicks, and vice versa for slowly rotating cores. Then, large kicks would be a feature of stars that retained deep convective envelopes long enough to allow a strong magnetic torque, generated by differential rotation between the core and the envelope, to spin the core down to the very slow rotation rate of the envelope. A low kick velocity imparted to the second (younger) neutron star ($< 50 \text{ km s}^{-1}$) was inferred from the analysis of large-eccentricity binary pulsar PSR J1811–1736 [64]. The large orbital period of this binary pulsar (18.8 d) then may suggest an evolutionary scenario with inefficient (if any) common envelope stage [79], i.e. the absence of deep convective shell in the supernova progenitor (a He-star). This conclusion can be regarded as supportive to ideas put forward in [312].

In principle, it is possible to assume some kick velocity during BH formation as well [227, 112, 328, 330, 285, 22, 469]. The similarity of NS and BH distribution in the Galaxy suggesting BH kicks was noted in [174]. A recent analysis of the space velocity of some BH binary systems [452]

put an upper limit on the BH kick velocity of less than $\sim 200 \text{ km s}^{-1}$. However, no kick seems to be required to explain the formation of other BH candidates (Cyg X-1, X-Nova Sco, etc.) [282].

To summarize, the kick velocity remains one of the important unknown parameters of binary evolution with NSs and BHs. Further constraining this parameter from various observations and theoretical understanding of possible asymmetry of core-collapse supernovae seem to be of paramount importance for the formation and evolution of close compact binaries.

3.4.1 Effect of the kick velocity on the evolution of a binary system

The collapse of a star to a BH, or its explosion leading to the formation of a NS, are normally considered as instantaneous. This assumption is well justified in binary systems, since typical orbital velocities before the explosion do not exceed a few hundred km/s, while most of the mass is expelled with velocities about several thousand km/s. The exploding star M_1 leaves the remnant M_c , and the binary loses a portion of its mass: $\Delta M = M_1 - M_c$. The relative velocity of stars before the event is

$$V_i = \sqrt{G(M_1 + M_2)/a_i}. \quad (47)$$

Right after the event, the relative velocity is

$$\vec{V}_f = \vec{V}_i + \vec{w}. \quad (48)$$

Depending on the direction of the kick velocity vector \vec{w} , the absolute value of \vec{V}_f varies in the interval from the smallest $V_f = |V_i - w|$ to the largest $V_f = V_i + w$. The system gets disrupted if V_f satisfies the condition (see Section 3.3)

$$V_f \geq V_i \sqrt{\frac{2}{\chi}}, \quad (49)$$

where $\chi \equiv (M_1 + M_2)/(M_c + M_2)$.

Let us start from the limiting case when the mass loss is practically zero ($\Delta M = 0$, $\chi = 1$), while a non-zero kick velocity can still be present. This situation can be relevant to BH formation. It follows from Equation (49) that, for relatively small kicks, $w < (\sqrt{2} - 1)V_i$, the system always (independently of the direction of \vec{w}) remains bound, while for $w > (\sqrt{2} + 1)V_i$ the system always unbinds. By averaging over equally probable orientations of \vec{w} with a fixed amplitude w , one can show that in the particular case $w = V_i$ the system disrupts or survives with equal probabilities. If $V_f < V_i$, the semimajor axis of the system becomes smaller than the original binary separation, $a_f < a_i$ (see Equation (40)). This means that the system becomes more hard than before, i.e. it has a greater negative total energy than the original binary. If $V_i < V_f < \sqrt{2}V_i$, the system remains bound, but $a_f > a_i$. For small and moderate kicks $w \gtrsim V_i$, the probabilities for the system to become more or less bound are approximately equal.

In general, the binary system loses some fraction of its mass ΔM . In the absence of the kick, the system remains bound if $\Delta M < M/2$ and gets disrupted if $\Delta M \geq M/2$ (see Section 3.3). Clearly, a properly oriented kick velocity (directed against the vector \vec{V}_i) can keep the system bound, even if it would have been disrupted without the kick. And, on the other hand, an unfortunate direction of \vec{w} can disrupt the system, which otherwise would stay bound.

Consider, first, the case $\Delta M < M/2$. The parameter χ varies in the interval from 1 to 2, and the escape velocity V_e varies in the interval from $\sqrt{2}V_i$ to V_i . It follows from Equation (44) that the binary always remains bound if $w < V_e - V_i$, and always unbinds if $w > V_e + V_i$. This is a generalization of the formulae derived above for the limiting case $\Delta M = 0$. Obviously, for a given w , the probability for the system to disrupt or become softer increases when ΔM becomes larger. Now turn to the case $\Delta M > M/2$. The escape velocity of the compact star becomes $V_e < V_i$. The binary is always disrupted if the kick velocity is too large or too small: $w > V_i + V_e$ or $w < V_i - V_e$.

However, for all intermediate values of w , the system can remain bound, and sometimes even more bound than before, if the direction of \vec{w} happened to be approximately opposite to \vec{V}_i . A detailed calculation of the probabilities for the binary survival or disruption requires integration over the kick velocity distribution function $f(\vec{w})$ (see, e.g., [43]).

3.5 Common envelope stage

This is a very important stage in the evolution of binaries of all masses. In different contexts, the evolution of binary components in an envelope engulfing both of them was considered, for instance, in [344, 384, 6], but the importance of common envelopes was really recognized after Paczyński [301] and Ostriker [298] applied them to explain the formation of cataclysmic variables and massive X-ray binaries, respectively. A detailed review of problems related to common envelopes may be found, e.g., in [158, 393].

Generally, common envelopes form in binary systems where the mass transfer from the mass-losing star is high, and the companion cannot accommodate all the accreting matter. The common envelope stage appears unavoidable on observational grounds. The evidence for a dramatic orbital angular momentum decrease in some preceding evolutionary stage follows from observations of certain types of close binary stars. They include cataclysmic variables, in which a white dwarf accretes matter from a small red dwarf main-sequence companion, planetary nebulae with double cores, low-mass X-ray binaries and X-ray transients (neutron stars and black holes accreting matter from low-mass main-sequence dwarfs). The radii of progenitors of compact stars in these binaries typically should have been 100–1000 solar radii, that is, much larger than the currently observed binary separations. This testifies of some dramatic reduction of the orbital momentum in the earlier stages of evolution and eventual removal of the common envelope. Additional indirect evidence for reality of the common envelope stage in the typical pre-cataclysmic binary V471 Tau has recently been obtained from X-ray Chandra observations [85] showing anomalous C/N contamination of the K-dwarf companion. Recent studies also indicate that many planetary nebulae are actually binaries, which may suggest that most of them result from common envelope interaction [474, 73].

Exact criteria for the formation of a common envelope are absent. However, a high rate of mass overflow onto a compact star from a normal star is always expected when the normal star goes off the main sequence and develops a deep convective envelope. The physical reason for this is that convection tends to make entropy constant along the radius, so the radial structure of convective stellar envelopes is well described by a polytrope (i.e. the equation of state can be written as $P = K\rho^{1+1/n}$) with an index $n = 3/2$. The polytropic approximation with $n = 3/2$ is also valid for degenerate white dwarfs with masses not too close to the Chandrasekhar limit. For a star in hydrostatic equilibrium, this results in the well known inverse mass-radius relation, $R \propto M^{-1/3}$, first measured for white dwarfs. Removing mass from a star with a negative power of the mass-radius relation increases its radius. On the other hand, the Roche lobe of the more massive star should shrink in response to the conservative mass exchange between the components. This further increases the mass loss rate from the Roche-lobe filling star leading to a continuation of an unstable mass loss and eventual formation of a common envelope. The critical mass ratio for the unstable Roche lobe overflow depends on specifics of the stellar structure and mass ratio of components; typically, mass loss is unstable for stars with convective envelopes, stars with radiative envelopes if $q \gtrsim 2$, and white dwarfs if $q \gtrsim 2/3$.

As other examples for the formation of a common envelope one may consider, for instance, direct penetration of a compact star into the dense outer layers of the companion, which can happen as a result of the Darwin tidal orbital instability in binaries [65, 12]; it is possible that a compact remnant of a supernova explosion with appropriately directed kick velocity finds itself in an elliptic orbit whose minimum periastron distance $a_f(1 - e)$ is smaller than the stellar radius of the companion; a common envelope enshrouding both components of a binary may form due to

unstable thermonuclear burning in the surface layers of an accreting WD.

The common envelope stage is, usually, treated in the following simplified way [444, 72]. The orbital evolution of the compact star m inside the envelope of the normal star M_1 is driven by the dynamical friction drag. This leads to a gradual spiral-in process of the compact star. The released orbital energy ΔE_{orb} , or a fraction of it, can become numerically equal to the binding energy E_{bind} of the envelope with the rest of the binary system. It is generally assumed that the orbital energy of the binary is used to expel the envelope of the donor with an efficiency α_{ce} :

$$E_{\text{bind}} = \alpha_{\text{ce}} \Delta E_{\text{orb}},$$

where E_{bind} is the total binding energy of the envelope and ΔE_{orb} is the orbital energy released in the spiral-in. What remains of the normal star M_1 is its stellar core M_c . The above energy condition reads

$$\frac{GM_1(M_1 - M_c)}{\lambda R_L} = \alpha_{\text{ce}} \left(\frac{GmM_c}{2a_f} - \frac{GM_1m}{2a_i} \right), \quad (50)$$

where a_i and a_f are the initial and the final orbital separations, and λ is a numerical coefficient that depends on the structure of the donor's envelope. R_L is the Roche lobe radius of the normal star that can be approximated as [89]

$$\frac{R_L}{a_i} = \frac{0.49}{0.6 + q^{2/3} \ln(1 + q^{1/3})} \quad (51)$$

and $q \equiv M_1/m$. From Equation (50) one derives

$$\frac{a_f}{a_i} = \frac{M_c}{M_1} \left(1 + \frac{2a_i}{\lambda \alpha_{\text{ce}} R_L} \frac{M_1 - M_c}{m} \right)^{-1} \lesssim \frac{M_c}{M_1} \frac{M_c}{\Delta M}, \quad (52)$$

where $\Delta M = M_1 - M_c$ is the mass of the ejected envelope. The mass M_c of a helium core of a massive star may be approximated as [415]

$$M_{\text{He}} \approx 0.1(M_1/M_\odot)^{1.4}, \quad (53)$$

so the orbital separation during the common envelope stage may decrease as much as by factor 100 and more.

The above formalism for the common envelope stage depends in fact on the product of two parameters: λ , which is the measure of the binding energy of the envelope to the core prior to mass transfer in a binary system, and α_{ce} , which is the common envelope efficiency itself. Numerical calculations of evolved giant stars with masses $3-10 M_\odot$ [80] showed that the value of the λ -parameter is typically between 0.2 and 0.8; however, it can be as high as 5 on the asymptotic giant branch. For more massive primaries ($> 20 M_\odot$), which are appropriate for the formation of BH binaries, the λ -parameter was found to depend on the mass of the star and vary within a wide range 0.01–0.5 [320]. Some hydrodynamical simulations [341] indicated that $\alpha_{\text{ce}} \simeq 1$, while in others [367] a wider range for values of α_{ce} was obtained. There are debates in the literature as to should additional sources of energy (e.g., ionization energy in the envelope [134]) should be included in the ejection criterion of common envelopes [381].

There is another approach, different from the standard Webbink formalism, which is used to estimate the common envelope efficiency α_{ce} . In the case of systems with at least one white-dwarf component one can try to reconstruct the evolution of double compact binaries with known masses of both components, since there is a unique relation between the mass of a white dwarf and the radius of its red giant progenitor. Close binary white dwarfs should definitely result from the spiral-in phase in the common envelope that appears inevitable during the second mass transfer (i.e. from the red giant to the white dwarf remnant of the original primary in the system). Such

an analysis [284], extended in [283, 431], suggests that the standard energy prescription for the treatment of the common envelope stage cannot be applied to the first mass transfer episode. Instead, the authors proposed to apply the so-called γ -formalism for the common envelope, in which not the energy but the angular momentum J is balanced and conservation of energy is implicitly implied:

$$\frac{\delta J}{J} = \gamma \frac{\Delta M}{M_{\text{tot}}}, \quad (54)$$

where ΔM is the mass of the common envelope, M_{tot} is the total mass of the binary system before the common envelope, and γ is a numerical coefficient. For all binary systems considered, the parameter γ was found to be within the range 1.1–4, with the mean value around 1.5. (Notice that $\gamma = 1$ corresponds to loss of the angular momentum through stellar wind, as considered above, which always increases the orbital separation of the binary.) The applicability of this algorithm should be investigated further.

Note also, that formulations of the common envelope equation different from Equation (50) are met in the literature (see, e.g., [416, 158]); a_f/a_i similar to the values produced by Equation (50) are then obtained for different $\alpha_{\text{ce}}\lambda$ values.

4 Evolutionary Scenario for Compact Binaries with Neutron Star or Black Hole Components

4.1 Compact binaries with NSs

Compact binaries with NS and BH components are descendants of initially massive binaries with $M_1 \gtrsim (8-10)M_\odot$. The evolutionary scenario of massive binaries was elaborated shortly after the discovery of binary X-ray sources [415, 408, 427] and is depicted in Figure 4.

This scenario is fully confirmed by more than 30-years' history of astronomical observations and is now considered as standard. A massive X-ray binary is an inevitable stage preceding the formation of a double compact system after the second supernova explosion of the helium-rich companion in such a stellar system. The formation scenario for binary pulsars proposed immediately after the discovery of PSR 1913+16 [106, 253] has also been tested by subsequent observations of binary pulsars. In fact, the scenario for binary pulsars was proposed even earlier in [415], but because no binary pulsars were known at that time, it was suggested that all pairs of NS are disrupted at the second NS formation.

It is convenient to separate the evolution of a massive binary into several stages according to the physical state of the binary components, including phases of mass exchange between them. The simplest evolutionary scenario can be schematically described as follows (see Figure 4).

1. Initially, two high-mass OB main-sequence stars are separated and are inside their Roche lobes. Tidal interaction is very effective so that a possible initial eccentricity vanishes before the primary star M_1 fills its Roche lobe. The duration of this stage is determined by the hydrogen burning time of the primary and typically is several million years (for massive main-sequence stars, the time of core hydrogen burning is $t_{\text{nucl}} \propto M^{-2}$). The star burns out hydrogen in its central parts, so that a dense central helium core with a mass $M_{\text{He}} \simeq 0.1(M/M_\odot)^{1.4}$ forms by the time when the star leaves the main sequence. The expected number of such binaries in the Galaxy is around 10^4 .
2. After core hydrogen exhaustion, the primary leaves the main sequence and starts to expand rapidly. When its radius approaches the Roche lobe (see Equation (51)), mass transfer onto the secondary, less massive star which still resides on the main sequence begins. The mass-transfer rate can be crudely estimated as $\dot{M} \sim M_1/\tau_{\text{KH}}$, where $\tau_{\text{KH}} = GM_1^2/R_1L_1$ is the primary's thermal time scale.

The mass transfer ends when most of the primary's hydrogen envelope is transferred onto the secondary, so a naked helium core is left. This core can be observed as a Wolf-Rayet (WR) star with intense stellar wind if its mass exceeds $(7-8)M_\odot$ [293, 94, 95].

While the mass of the primary star reduces, the mass of the secondary star increases, since the mass transfer at this stage is thought to be quasi-conservative. For not too massive main-sequence stars, $M \lesssim 40M_\odot$, no significant stellar wind mass loss occurs which could, otherwise, remove too much matter from the binary, thereby increasing binary separation. The secondary star acquires large angular momentum due to the infalling material, so that its outer envelope can be spun up to an angular velocity close to the limiting (Kepler orbit) value. Such massive rapidly rotating stars are observed as Be-stars. During the conservative stage of mass transfer, the semimajor axis of the orbit first decreases, reaches a minimum when the masses of the binary components become equal to each other, and then increases. This behavior is dictated by the angular momentum conservation law (31). After the completion of the conservative mass transfer, the initially more massive star becomes less massive than its initially lighter companion.

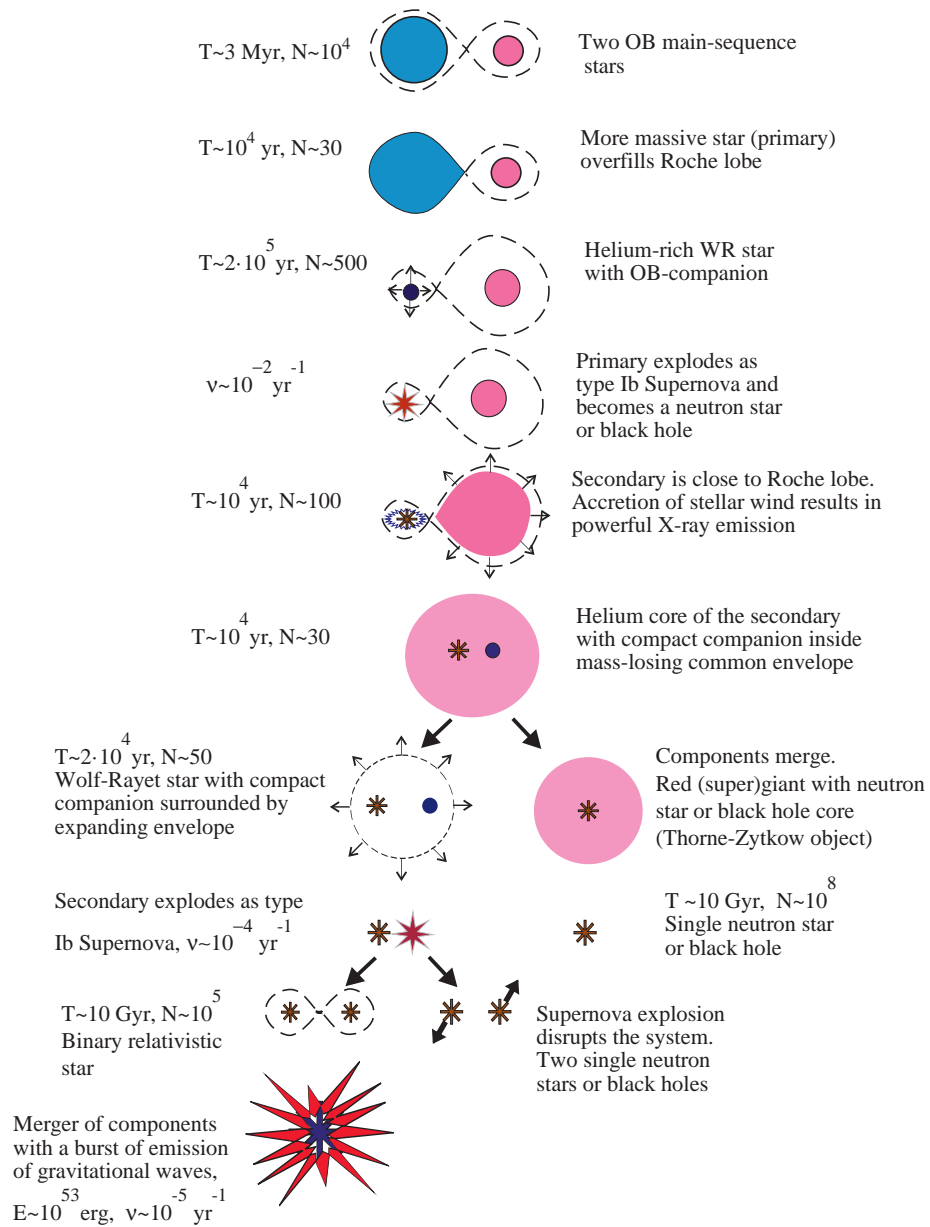


Figure 4: *Evolutionary scenario for the formation of neutron stars or black holes in close binaries.*

For the typical parameters the duration of the first RLOF is rather short, of the order of 10^4 yr, so only several dozens of such binaries are expected to be in the Galaxy.

3. The duration of the WR stage is about several 10^5 yr, so the Galactic number of such binaries should be several hundreds.
4. At the end of the thermonuclear evolution, the WR star explodes as Ib (or Ic) supernova to become a NS or BH. The inferred Galactic type Ib SN rate is around 10^{-2} per year, at least half of them should be in binaries. At this stage the disruption of the binary is possible (e.g., if the mass lost during the symmetric SN explosion exceeds 50% of the total mass of the pre-SN binary, or even smaller in the presence of the kick velocity; see Section 3.3 above). Some runaway Galactic OB-stars must have been formed in this way.
5. If the system survives the first SN explosion, a rapidly rotating Be star in pair with a young NS appears. Orbital evolution following the SN explosion is described above by Equations (40–45). The orbital eccentricity after the SN explosion is high, so enhanced accretion onto the NS occurs at the periastron passages. Most of about 100 Galactic Be/X-ray binaries [335] are formed in this way. The duration of this stage depends on the binary parameters, but in all cases it is limited by the time left for the (now more massive) secondary to burn hydrogen in its core.

An important parameter of NS evolution is the surface magnetic field strength. In binary systems, magnetic field, in combination with NS spin period and accretion rate onto the NS surface, determines the observational manifestation of the neutron star (see [219] for more detail). Accretion of matter onto the NS can reduce the surface magnetic field and spin-up the NS rotation (pulsar recycling) [37, 356, 357, 36].

6. The secondary expands to engulf the NS. The common envelope stage begins and after $\sim 10^3$ yr ends up with the formation of a WR star with a compact companion surrounded by an expanding envelope (Cyg X-3 may serve as an example), or the NS merges with the helium core during the common envelope to form a still hypothetical Thorne–Żytkow (TZ) object. The fate of TZ stars remains unclear (see [17] for the recent study). Single (possibly, massive) NS or BH should descend from them.

A note should be made concerning the phase when a common envelope engulfs the first-formed NS and the core of the secondary. Colgate [60] and Zel’dovich et al. [475] have shown that hyper-Eddington accretion onto a neutron star is possible if the gravitational energy released in accretion is lost by neutrinos. Chevalier [55] suggested that this may be the case for the accretion in common envelopes. Since the accretion rates in this case may be as high as $\sim 0.1 M_{\odot} \text{ yr}^{-1}$, the NS may collapse into a BH inside the common envelope. An essential caveat is that the accretion in the hyper-Eddington regime may be prevented by the angular momentum of the captured matter. The magnetic field of the NS may also be a complication. The possibility of hyper-critical accretion still has to be studied. Nevertheless, implications of this hypothesis for different types of relativistic binaries were explored in great detail by H. Bethe and G. Brown and their coauthors (see, e.g., [44] and references therein). Also, the possibility of hyper-Eddington accretion was included in several population synthesis studies with evident result of diminishing the population of NS + NS binaries in favour of neutron stars in pairs with low-mass black holes (see, e.g., [328, 22]).

7. The secondary WR ultimately explodes as a type Ib supernova leaving behind a double NS binary, or the system is disrupted to form two single high-velocity NSs or BHs. Even for a symmetric SN explosion the disruption of binaries after the second SN explosion could result in the observed high average velocities of radiopulsars (see Section 3.4 above). In the

surviving close binary NS system, the older NS is expected to have faster rotation velocity (and possibly higher mass) than the younger one because of the recycling at the preceding accretion stage. The subsequent orbital evolution of such double NS systems is entirely due to GW emission (see Section 3.1.4) and ultimately leads to the coalescence of the components.

Detailed studies of possible evolutionary channels which produce merging binary NS can be found in the literature (see, e.g., [419, 420, 227, 328, 13, 449, 22, 169, 81, 453]).

We emphasize that this scenario applies only to initially massive binaries. There exists also a population of NSs accompanied by low-mass [$\sim (1-2) M_{\odot}$] companions. A scenario similar to the one presented in Figure 4 may be sketched for them too, with the difference that the secondary component stably transfers mass onto the companion (see, e.g., [165, 183, 184, 410]). This scenario is similar to the one for low- and intermediate-mass binaries considered in Section 7, with the WD replaced by a NS or a BH. Compact low-mass binaries with NSs may be dynamically formed in dense stellar environments, for example in globular clusters. The dynamical evolution of binaries in globular clusters is beyond the scope of this review; see [26] and [36] for more detail and further references.

4.2 Black hole formation parameters

So far, we have considered the formation of NSs and binaries with NSs. It is believed that very massive stars end up their evolution with the formation of stellar mass black holes. We will discuss now their formation.

In the analysis of BH formation, new important parameters appear. The first one is the threshold mass M_{cr} beginning from which a main-sequence star, after the completion of its nuclear evolution, can collapse into a BH. This mass is not well known; different authors suggest different values: van den Heuvel and Habets [428] – $40 M_{\odot}$; Woosley et al. [458] – $60 M_{\odot}$; Portegies Zwart, Verbunt, and Ergma [327] – more than $20 M_{\odot}$. A simple physical argument usually put forward in the literature is that the mantle of the main-sequence star with $M > M_{\text{cr}} \approx 30 M_{\odot}$ before the collapse has a binding energy well above 10^{51} erg (the typical supernova energy observed), so that the supernova shock is not strong enough to expel the mantle [110, 111].

The upper mass limit for BH formation (with the caveat that the role of magnetic-field effects is not considered) is, predominantly, a function of stellar-wind mass loss in the core-hydrogen, hydrogen-shell, and core-helium burning stages. For a specific combination of winds in different evolutionary stages and assumptions on metallicity it is possible to find the types of stellar remnants as a function of initial mass (see, for instance [138]). Since stellar winds are mass (or luminosity) and metallicity-dependent, a peculiar consequence of mass-loss implementation in the latter study is that for $Z \simeq Z_{\odot}$ the mass-range of precursors of black holes is constrained to $M \approx (25-60) M_{\odot}$, while more massive stars form NSs because of heavy mass loss. The recent discovery of the possible magnetar in the young stellar cluster Westerlund 1 [263] hints to the reality of such a scenario. Note, however, that the estimates of \dot{M} are rather uncertain, especially for the most massive stars, mainly because of clumping in the winds (see, e.g., [201, 69, 129]). Current reassessment of the role of clumping generally results in the reduction of previous mass-loss estimates. Other factors that have to be taken into account in the estimates of the masses of progenitors of BHs are rotation and magnetic fields.

The second parameter is the mass M_{BH} of the nascent BH. There are various studies as for what the mass of the BH should be (see, e.g., [403, 32, 110, 113]). In some papers a typical BH mass was found to be not much higher than the upper limit for the NS mass (Oppenheimer–Volkoff limit $\sim (1.6-2.5) M_{\odot}$, depending on the unknown equation of state for NS matter) even if the fallback accretion onto the supernova remnant is allowed [403]. Modern measurements of black hole masses in binaries suggest a broad range of BH masses of the order of $4-17 M_{\odot}$ [296, 254, 345]. A continuous range of BH masses up to $10-15 M_{\odot}$ was derived in calculations [113]. Since present

day calculations are still unable to reproduce self-consistently even the supernova explosion, in the further discussion we have parameterized the BH mass M_{BH} by the fraction of the pre-supernova mass M_* that collapses into the BH: $k_{\text{BH}} = M_{\text{BH}}/M_*$. In fact, the pre-supernova mass M_* is directly related to M_{cr} , but the form of this relationship is somewhat different in different scenarios for massive star evolution, mainly because of different mass-loss prescriptions. According to our parameterization, the minimal BH mass can be $M_{\text{BH}}^{\text{min}} = k_{\text{BH}}M_*$, where M_* itself depends on M_{cr} . The parameter k_{BH} can vary in a wide range.

The third parameter, similar to the case of NS formation, is the possible kick velocity \mathbf{w}_{BH} imparted to the newly formed BH (see the end of Section 3.4). In general, one expects that the BH should acquire a smaller kick velocity than a NS, as black holes are more massive than neutron stars. A possible relation (as adopted, e.g., in calculations [227]) reads

$$\frac{w_{\text{BH}}}{w_{\text{NS}}} = \frac{M_* - M_{\text{BH}}}{M_* - M_{\text{OV}}} = \frac{1 - k_{\text{BH}}}{1 - M_{\text{OV}}/M_*}, \quad (55)$$

where $M_{\text{OV}} = 2.5 M_{\odot}$ is the maximum NS mass. When M_{BH} is close to M_{OV} , the ratio $w_{\text{BH}}/w_{\text{NS}}$ approaches 1, and the low-mass black holes acquire kick velocities similar to those of neutron stars. When M_{BH} is significantly larger than M_{OV} , the parameter $k_{\text{BH}} = 1$, and the BH kick velocity becomes vanishingly small. The allowance for a quite moderate w_{BH} can increase the coalescence rate of binary BH [227].

The possible kick velocity imparted to newly born black holes makes the orbits of survived systems highly eccentric. It is important to stress that some fraction of such binary BH can retain their large eccentricities up to the late stages of their coalescence. This signature should be reflected in their emitted waveforms and should be modeled in templates.

Asymmetric explosions accompanied by a kick change the space orientation of the orbital angular momentum. On the other hand, the star's spin axis remains fixed (unless the kick was off-center). As a result, some distribution of the angles between the BH spins and the orbital angular momentum (denoted by J) will be established [330]. It is interesting that even for small kicks of a few tens of km/s an appreciable fraction (30–50%) of the merging binary BH can have $\cos J < 0$. This means that in these binaries the orbital angular momentum vector is oriented almost oppositely to the black hole spins. This is one more signature of imparted kicks that can be tested observationally. These effects are also discussed in [179].

5 Formation of Double Compact Binaries

5.1 Analytical estimates

A rough estimate of the formation rate of double compact binaries can be obtained ignoring many details of binary evolution. To do this, we shall use the observed initial distribution of binary orbital parameters and assume the simplest conservative mass transfer ($M_1 + M_2 = \text{const}$) without kick velocity imparted to the nascent compact stellar remnants during SN explosions.

Initial binary distributions. From observations of spectroscopic binaries it is possible to derive the formation rate of binary stars with initial masses M_1 , M_2 (with mass ratio $q = M_2/M_1 \leq 1$), orbital semimajor axis A , and eccentricity e . According to [326], the present birth rate of binaries in our Galaxy can be written in factorized form as

$$\frac{dN}{dA dM_1 dq dt} \approx 0.087 \left(\frac{A}{R_\odot} \right)^{-1} \left(\frac{M_1}{M_\odot} \right)^{-2.5} f(q), \quad (56)$$

where $f(q)$ is a poorly constrained distribution over the initial mass ratio of binary components.

One usually assumes a mass ratio distribution law in the form $f(q) \sim q^{-\alpha_q}$ where α_q is a free parameter; another often used form of the q -distribution was suggested by Kuiper [202]:

$$f(q) = 2/(1+q)^2.$$

The range of A is $10 \leq A/R_\odot \leq 10^6$. In deriving the above Equation (56), the authors of [326] took into account selection effects to convert the ‘‘observed’’ distribution of stars into the true one. An almost flat logarithmic distribution of semimajor axes was also found in [4]. Integration of Equation (56) yields one binary system with $M_1 \gtrsim 0.8M_\odot$ and $10R_\odot < A < 10^6R_\odot$ per year in the Galaxy, which is in reasonable agreement with the Galactic star formation rate estimated by various methods; the present-day star formation rate is about several M_\odot per year (see, for example, [255, 401]).

Constraints from conservative evolution. To form a NS at the end of thermonuclear evolution, the primary mass should be at least $10M_\odot$. Equation (56) says that the formation rate of such binaries is about 1 per 50 years. We shall restrict ourselves by considering only close binaries, in which mass transfer onto the secondary is possible. This narrows the binary separation interval to $10 - 1000R_\odot$ (see Figure 1); the birth rate of close massive ($M_1 > 10M_\odot$) binaries is thus $1/50 \times 2/5 \text{ yr}^{-1} = 1/125 \text{ yr}^{-1}$. The mass ratio q should not be very small to make the formation of the second NS possible. The lower limit for q is derived from the condition that after the first mass transfer stage, during which the mass of the secondary increases, $M_2 + \Delta M \geq 10M_\odot$. Here $\Delta M = M_1 - M_{\text{He}}$ and the mass of the helium core left after the first mass transfer is $M_{\text{He}} \approx 0.1(M_1/M_\odot)^{1.4}$. This yields

$$m_2 + (m_1 - 0.1m_1^{1.4}) > 10,$$

where we used the notation $m = M/M_\odot$, or in terms of q :

$$q \geq 10/m_1 + 0.1m_1^{0.4} - 1. \quad (57)$$

An upper limit for the mass ratio is obtained from the requirement that the binary system remains bound after the sudden mass loss in the second supernova explosion⁵. From Equation (45) we obtain

$$\frac{0.1[m_2 + (m_1 - 0.1m_1^{1.4})]^{1.4} - 1.4}{2.8} < 1,$$

⁵For the first supernova explosion without kick this is always satisfied.

or in terms of q :

$$q \leq 14.4/m_1 + 0.1m_1^{0.4} - 1. \quad (58)$$

Inserting $m_1 = 10$ in the above two equations yields the appropriate mass ratio range $0.25 < q < 0.69$, i.e. 20% of the binaries for Kuiper’s mass ratio distribution. So we conclude that the birth rate of binaries which potentially can produce double NS system is $\lesssim 0.2 \times 1/125 \text{ yr}^{-1} \simeq 1/600 \text{ yr}^{-1}$.

Of course, this is a very crude upper limit – we have not taken into account the evolution of the binary separation, ignored initial binary eccentricities, non-conservative mass loss, etc. However, it is not easy to treat all these factors without additional knowledge of numerous details and parameters of binary evolution (such as the physical state of the star at the moment of the Roche lobe overflow, the common envelope efficiency, etc.). All these factors should decrease the formation rate of double NS. The coalescence rate of compact binaries (which is ultimately of interest for us) will be even smaller – for the compact binary to merge within the Hubble time, the binary separation after the second supernova explosion should be less than $\sim 100 R_\odot$ (orbital periods shorter than ~ 40 d) for arbitrary high orbital eccentricity e (see Figure 3). The model-dependent distribution of NS kick velocities provides another strong complication. We also stress that this upper limit was obtained assuming a constant Galactic star-formation rate and normalization of the binary formation by Equation (56).

Further (semi-)analytical investigations of the parameter space of binaries leading to the formation of coalescing binary NSs are still possible but technically very difficult, and we shall not reproduce them here. The detailed semi-analytical approach to the problem of formation of NSs in binaries and evolution of compact binaries has been developed by Tutukov and Yungelson [419, 420].

5.2 Population synthesis results

A distinct approach to the analysis of binary star evolution is based on the population synthesis method – a Monte-Carlo simulation of the evolution of a sample of binaries with different initial parameters. This approach was first applied to model various observational manifestations of magnetized NSs in massive binary systems [195, 196, 78] and generalized to binary systems of arbitrary mass in [224] (The Scenario Machine code). To achieve a sufficient statistical significance, such simulations usually involve a large number of binaries, typically of the order of a million. The total number of stars in the Galaxy is still four orders of magnitude larger, so this approach cannot guarantee that rare stages of the binary evolution will be adequately reproduced⁶.

Presently, there are several population synthesis codes used for massive binary system studies, which take into account with different degree of completeness various aspects of binary stellar evolution (e.g., the codes by Portegies Zwart et al. [328, 469], Bethe and Brown [32], Hurley, Tout, and Pols [154], Belczynski et al. [23], Yungelson and Tutukov [423]). A review of applications of the population synthesis method to various types of astrophysical sources and further references can be found in [325, 467]. Some results of population synthesis calculations of compact binary mergers carried out by different groups are presented in Table 4.

Actually, the authors of the studies mentioned in Table 4 make their simulations for a range of parameters. We list in the table the rates for the models which the authors themselves consider as “standard” or “preferred” or “most probable”. Generally, for the NS + NS merger rate Table 4 shows the scatter within a factor ~ 4 , which may be considered quite reasonable, having in mind the uncertainties in input parameters. There are two clear outliers, [420] and [437]. The high rate in [420] is due to the assumption that kicks to nascent neutron stars are absent. The low rate in [437] is due to the fact that these authors apply in the common envelope equation an evolutionary-stage-dependent structural constant λ . Their range for λ is 0.006–0.4, to be compared with the

⁶Instead of Monte-Carlo simulations one may use a sufficiently dense grid in the 3D space of binary parameters and integrate over this grid (see, e.g., [423] and references therein).

Authors	Ref.	NS + NS [yr ⁻¹]	NS + BH [yr ⁻¹]	BH + BH [yr ⁻¹]
Tutukov and Yungelson (1993)	[420]	3×10^{-4}	2×10^{-5}	1×10^{-6}
Lipunov et al. (1997)	[226]	3×10^{-5}	2×10^{-6}	3×10^{-7}
Portegies Zwart and Yungelson (1998)	[328]	2×10^{-5}	10^{-6}	
Nelemans et al. (2001)	[285]	2×10^{-5}	4×10^{-6}	
Voss and Tauris (2003)	[437]	2×10^{-6}	6×10^{-7}	10^{-5}
O’Shaughnessy et al. (2005)	[297]	7×10^{-6}	1×10^{-6}	1×10^{-6}
de Freitas Pacheco et al. (2006)	[71]	2×10^{-5}		

Table 4: *Examples of the estimates for Galactic merger rates of relativistic binaries calculated under different assumptions on the parameters entering population synthesis.*

“standard” $\lambda = 0.5$ applied in most of the other studies. A low λ favours mergers in the first critical lobe overflow episode and later mergers of the first-born neutron stars with their non-relativistic companions⁷. A considerable scatter in the rates of mergers of systems with BH companions is due, mainly, to uncertainties in stellar wind mass loss for the most massive stars. For instance, the implementation of winds in the code used in [328, 285] resulted in the absence of merging BH + BH systems, while a rather low \dot{M} assumed in [437] produced a high merger rate of BH + BH systems.

A word of caution should be said here. It is hardly possible to trace a detailed evolution of each binary, so one usually invokes the approximate approach to describe the change of evolutionary stages of the binary components (the so-called evolutionary track), their interaction, effects of supernovae, etc. Thus, fundamental uncertainties of stellar evolution mentioned above are complemented with (i) uncertainties of the scenario and (ii) uncertainties in the normalization of the calculations to the real galaxy (such as the fraction of binaries among allstars, the star formation history, etc.). The intrinsic uncertainties in the population synthesis results (for example, in the computed event rates of binary mergers etc.) are in the best case not less than of the order of factor two or three. This should always be born in mind when using the population synthesis calculations. However, we emphasize again the fact that the double NS merger rate, as inferred from binary pulsar statistics with account for the double pulsar observations [48, 180], is very close to the population syntheses estimates with a kick of about $(250 - 300) \text{ km s}^{-1}$.

⁷Note that similar low values of λ for 20 to $50 M_{\odot}$ stars were obtained also by [320]. If confirmed, these results may have major impact on the estimates of merger rates for relativistic binaries.

6 Detection Rates

The detection of a gravitational wave signal from merging close binaries is characterized by the signal-to-noise ratio S/N , which depends on the binary masses, the distance to the binary, the frequency, and the noise characteristics. A pedagogical derivation of the signal-to-noise ratio and its discussion for different detectors is given, for example, in Section 8 of the review [122].

In this section we focus on two particular points: the plausible enhancement of the detection of merging binary black holes with respect to binary neutron stars and the way how absolute detection rates of binary mergings can be calculated.

6.1 Enhancement of the detection rate for binary BH mergers

Coalescing binaries emit gravitational wave signals with a well known time-dependence (waveform) (see Section 3.1 above). This allows one to use the technique of matched filtering [399]. The signal-to-noise ratio S/N for a particular detector, which is characterized by the dimensionless noise rms amplitude h_{rms} at a given frequency f , depends mostly on the “chirp” mass of the binary system $\mathcal{M} = (M_1 + M_2)^{-1/5} (M_1 M_2)^{3/5}$ and its distance r . Here, we will use the simplified version for S/N ([399]; see also [105]):

$$\frac{S}{N} = 3^{-1/2} \pi^{-2/3} \frac{G^{5/6}}{c^{3/2}} \frac{\mathcal{M}^{5/6}}{r} \frac{f^{-1/6}}{h_{\text{rms}}(f)}. \quad (59)$$

At a fixed level of S/N , the detection volume is proportional to r^3 and therefore it is proportional to $\mathcal{M}^{5/2}$. The detection rate \mathcal{D} for binaries of a given class (NS + NS, NS + BH or BH + BH) is the product of their coalescence rate \mathcal{R}_V with the detector’s registration volume $\propto \mathcal{M}^{5/2}$ for these binaries.

It is seen from Table 4 that the model Galactic rate \mathcal{R}_G of NS + NS coalescences is typically higher than the rate of NS + BH and BH + BH coalescences. However, the BH mass can be significantly larger than the NS mass. So a binary involving one or two black holes, placed at the same distance as a NS + NS binary, produces a significantly larger amplitude of gravitational waves. With the given sensitivity of the detector (fixed S/N ratio), a BH + BH binary can be seen at a greater distance than a NS + NS binary. Hence, the registration volume for such bright binaries is significantly larger than the registration volume for relatively weak binaries. The detection rate of a given detector depends on the interplay between the coalescence rate and the detector’s response to the sources of one or another kind.

If we assign some characteristic (mean) chirp mass to different types of double NS and BH systems, the expected ratio of their detection rates by a given detector is

$$\frac{\mathcal{D}_{\text{BH}}}{\mathcal{D}_{\text{NS}}} = \frac{\mathcal{R}_{\text{BH}}}{\mathcal{R}_{\text{NS}}} \left(\frac{\mathcal{M}_{\text{BH}}}{\mathcal{M}_{\text{NS}}} \right)^{5/2}, \quad (60)$$

where \mathcal{D}_{BH} and \mathcal{D}_{NS} refer to BH + BH and NS + NS pairs, respectively. Here, we discuss the ratio of the detection rates, rather than their absolute values. The derivation of absolute values requires detailed evolutionary calculations, as we discussed above. Taking $\mathcal{M}_{\text{BH}} = 8.7 M_\odot$ (for $10 M_\odot + 10 M_\odot$) and $\mathcal{M}_{\text{NS}} = 1.22 M_\odot$ (for $1.4 M_\odot + 1.4 M_\odot$), Equation (60) yields

$$\frac{\mathcal{D}_{\text{BH}}}{\mathcal{D}_{\text{NS}}} \approx 140 \frac{\mathcal{R}_{\text{BH}}}{\mathcal{R}_{\text{NS}}}. \quad (61)$$

As $\frac{\mathcal{R}_{\text{BH}}}{\mathcal{R}_{\text{NS}}}$ is typically 0.1–0.01 (see Table 4), this relation suggests that the registration rate of BH mergers can be *higher* than that of NS mergers. This estimate is, of course, very rough, but it can serve as an indication of what one can expect from detailed calculations. We stress that the effect of an enhanced detection rate of BH binaries is independent of the desired S/N and other characteristics of the detector; it was discussed, for example, in [420, 226, 122].

6.2 Note on realistic calculation of the detection rates of binary mergings

Now we shall briefly discuss how the detection rates of binary mergings can be calculated for a given gravitational wave detector. For a secure detection, the S/N ratio is usually raised up to 7–8 to avoid false alarms over a period of a year (assuming Gaussian noise)⁸. This requirement determines the maximum distance from which an event can be detected by a given interferometer. The distance ranges of LIGO I/VIRGO (LIGO II) interferometers for relativistic binary inspirals are given in [70]: 20(300) Mpc for NS + NS ($1.4 M_{\odot} + 1.4 M_{\odot}$), 43(650) Mpc for NS + BH ($1.4 M_{\odot} + 10 M_{\odot}$). Note that the distances increase for a network of detectors.

To calculate a realistic detection rate of binary mergers the distribution of galaxies should be taken into account within the volume bounded by the distance range of the detector (see, for example, the earlier attempt to take into account only bright galaxies from Tully’s catalog of nearby galaxies in [221], and the use of LEDA database of galaxies to estimate the detection rate of supernova explosions [19]). However, not only the mass and type of a given galaxy, but also the star formation rate and, better, the history of the star formation rate in that galaxy (since the binary merger rate in galaxies strongly evolves with time [229]) are needed to estimate the expected detection rate \mathcal{D} . But this is a tremendous problem – even the sample of galaxies (mostly, dwarfs) within the Local Volume (< 8 Mpc) is only 70%–80% complete [185], and the number of new nearby galaxies continues to increase. So to assess the merger rate from a large volume based on the Galactic values, the best one can do at present appears to be using formulas like Equation (4) given earlier in Section 2.2. This, however, adds another factor two of uncertainty in the estimates. Clearly, a more accurate treatment of the transition from Galactic rates to larger volumes with an account of the galaxy distribution is required.

6.3 Going further

The most important future observations include:

1. Routine increase of the statistics of binary pulsars, especially with low radio luminosity. This will allow one to put stronger constraints on the NS + NS merger rate as directly inferred from the binary pulsars statistics. More indirectly, a larger sample of NS parameters in binary pulsars would be useful for constraining the range of parameters of scenarios of formation for double NSs and, hence, a better understanding their origin (see, for example, a recent attempt of such an analysis in [297]).
2. Discovery of a possible NS + BH binary. Measurements of its parameters would be crucial for models of formation and evolution of BHs in binary systems in general. Current estimates of the number of such binaries in the Galaxy, obtained by the population synthesis method, range from one per several thousand ordinary pulsars [228, 220] to much smaller values of about 0.1–1% of the number of double NSs in the Galactic disk [309].
3. Search for unusual observational manifestations of relativistic binaries (e.g., among some new radio transient sources like GCRT J1745-3009, firm identifications with some GRBs, etc.)
4. Improving the estimates of binary merger rate limits from data taken by GW detectors.

The most important theoretical issues include:

1. Stellar physics: post-helium burning evolution of massive stars, supernova explosion mechanism, masses of compact stars formed in the collapse, mechanism(s) of kick velocity imparted

⁸If a network of three detectors, such as two LIGOs and VIRGO, runs simultaneously, the S/N ratio in an individual detector should be $> 7/\sqrt{3} \approx 4$

to nascent compact remnants (neutron stars and black holes), stellar winds from hydrogen- and helium-rich stars.

2. Binary evolution: treatment of the common envelope stage, magnetic braking for low-mass binaries, observational constraints on the initial distributions of orbital parameters of binary stars (masses, semimajor axes, eccentricities).
3. Last but not least, it is very important to improve our knowledge of such “traditional” topics of stellar astronomy as the fraction of binary stars among the total population, distributions of binary stars over separations of components, and their mass ratios.

7 Formation of Short-Period Binaries with a White-Dwarf Components

Binary systems with white dwarf components that are interesting for general relativity and cosmology come in several flavours:

- Detached binary white dwarfs or “double degenerates” (DDs, we shall use both terms as synonyms below).
- Cataclysmic variables (CVs) – a class of semidetached binary stars containing a white dwarf and a companion star that is usually a red dwarf or a slightly evolved star, a subgiant.
- A subclass of the former systems in which the Roche lobe is filled by another white dwarf or low-mass partially degenerate helium star (AM CVn-type stars or “interacting double-degenerates”, IDD). They appear to be important LISA verification sources.
- Detached systems with a white dwarf accompanied by a low-mass nondegenerate helium star (SD + WD systems).
- Ultracompact X-ray binaries (UCXBs) where one of the components is a NS, while the Roche lobe overflowing component is a WD.

As Figure 2 shows, compact binary stars emit gravitational waves within the sensitivity limits for space-based detectors if their orbital periods are from ~ 20 s to 20,000 s. This means that in principle all AM CVn-stars, UCXBs, a considerable fraction of all CVs with measured orbital periods, and some DD and SD + WD systems would be observable in GWs in the absence of confusion noise and sufficient sensitivity of detectors.

Though general relativity predicted that binary stars have to be the sources of gravitational waves in the 1920s, this prediction became a matter of actual interest only with the discovery of the $P_{\text{orb}} \approx 81.5$ min cataclysmic variable WZ Sge by Kraft, Mathews, and Greenstein in 1962 [198], who immediately recognized the significance of short-period binary stars as testbeds for gravitational waves physics. Another impetus to the study of binaries as sources of gravitational wave radiation (GWR) was imparted by the discovery of the ultra-short period binary HZ 29 = AM CVn ($P_{\text{orb}} \approx 18$ min) by Smak in 1967 [379]. Smak [379] and Paczyński [300] speculated that the latter system is a close pair of white dwarfs, without specifying whether it is detached or semidetached. Faulkner et al. [100] inferred the status of AM CVn as a “double-white-dwarf semidetached” nova. AM CVn was later classified as a cataclysmic variable after flickering typical for CVs was found for AM CVn by Warner and Robinson [442]⁹ and became a prototype of a subclass of binaries¹⁰.

The origin of all above mentioned classes of short-period binaries was understood after the notion of common envelopes and the formalism for their treatment were suggested in 1970s (see Section 3.5). A spiral-in of components in common envelopes allowed to explain how white dwarfs – former cores of highly evolved stars with radii of $\sim 100 R_{\odot}$ – may acquire companions separated by $\sim R_{\odot}$ only (for pioneering work see [301, 443, 417, 418, 160, 444]). We recall, however, that most studies of the formation of compact objects through common envelopes are based on a simple formalism of comparison of binding energy of the envelope with the orbital energy of the binary, thought to be the sole source of energy for the loss of the envelope as described in Section 3.5. Though full-scale hydrodynamic calculations of a common-envelope evolution exist, for instance a

⁹Flickering is a fast intrinsic brightness scintillation occurring on time scales from seconds to minutes with amplitudes of 0.01–1 mag and suggesting an ongoing mass transfer.

¹⁰Remarkably, a clear-cut confirmation of the binary nature of AM CVn and the determination of its true period awaited for almost 25 years [281].

series of papers by Taam and coauthors published over more than two decades (see [368, 367] and references therein), the process is still very far from comprehension.

We recall also that the stability and timescale of mass-exchange in a binary depends on the mass ratio of components q , the structure of the envelope of Roche-lobe filling star, and possible stabilizing effects of mass and momentum loss from the system [413, 144, 470, 127, 135, 171, 102, 47]. For stars with radiative envelopes, to the first approximation, mass exchange is stable if $q \lesssim 1.2$; for $1.2 \lesssim q \lesssim 2$ it proceeds in the thermal time scale of the donor; for $q \gtrsim 2$ it proceeds in the dynamical time scale. Mass loss occurs on a dynamical time scale if the donor has a deep convective envelope or if it is degenerate and conditions for stable mass exchange are not satisfied. It is currently commonly accepted, despite a firm observational proof is lacking, that the distribution of binaries over q is even or rises to small q (see Section 5). Since typically the accretion rate is limited either by the rate that corresponds to the thermal time scale of the accretor or its Eddington accretion rate, both of which are lower than the mass-loss rate by the donor, the overwhelming majority ($\sim 90\%$) of close binaries pass in their evolution through one to four stages of a common envelope.

An “initial donor mass – donor radius at RLOF” diagram showing descendants of stars after mass-loss in close binaries is presented in Figure 1. We should remember here that solar metallicity stars with $M \lesssim 0.95 M_{\odot}$ do not evolve past the core-hydrogen burning stage in Hubble time.

Formation of compact binaries with WDs. A flowchart schematically presenting the typical scenario for formation of low-mass compact binaries with white-dwarf components and some endpoints of evolution is shown in Figure 5. Of course, not all possible scenarios are plotted, but only the most probable routes to SNe Ia and systems that may emit potentially detectable gravitational waves. For simplicity, we consider only the most general case when the first RLOF results in the formation of a common envelope.

The overwhelming majority of stars overflow their Roche lobes when they have He- or CO-cores. In stars with a mass below $(2.0-2.5)M_{\odot}$, helium cores are degenerate and if these stars overflow the Roche lobe prior to He-ignition, they produce helium white dwarfs. Binaries with non-degenerate He-core donors ($M \gtrsim (2.0-2.5)M_{\odot}$) first form a He-star + MS-star pair that may be observed as a subdwarf (sdB or sdO) star with MS companion. When the He-star completes its evolution, a pair harbouring a CO white dwarf and a MS-star appears.

If after the first common-envelope stage the orbital separation of the binary $a \simeq$ several R_{\odot} and the WD has a low-mass ($\lesssim 1.5 M_{\odot}$) MS companion the pair may evolve into contact during the hydrogen-burning stage or shortly after because of loss of angular momentum by a magnetically coupled stellar wind and/or GW radiation. If, additionally, the mass-ratio of components is favourable for stable mass transfer a cataclysmic variable may form. If the WD belongs to the CO-variety and accreted hydrogen burns at the surface of the WD stably, the WD may accumulate enough mass to explode as a type Ia supernova; the same may happen if in the recurrent outbursts less mass is ejected than accreted (the so-called SD scenario for SNe Ia originally suggested by Whelan and Iben [451]; see, e.g., [217, 470, 126, 132, 102, 133, 215] and references therein for later studies).

Some CV systems that burn hydrogen stably or are in the stage of residual hydrogen burning after an outburst may be also observed as supersoft X-ray sources (see, e.g., [426, 339, 175, 471, 177, 176]).

If the WD belongs to the ONe-variety, it may experience an AIC into a neutron star due to electron captures on Ne and Mg, and a low-mass X-ray binary may be formed.

The outcome of the evolution of a CV is not completely clear. It was hypothesized that the donor may be disrupted when its mass decreases below several hundredth of M_{\odot} [358]. Note, that for $q \lesssim 0.02$, that may be attained in Hubble time, and the conventional picture of mass exchange may become non-valid, since the circularization radius becomes greater than the outer radius of the disk. Matter flowing in from the companion circularises onto unstable orbits. At $q \approx 0.02$, matter is added at R_{circ} onto orbits that can become eccentric due to the 3:1 resonance. At $q \approx 0.005$ the

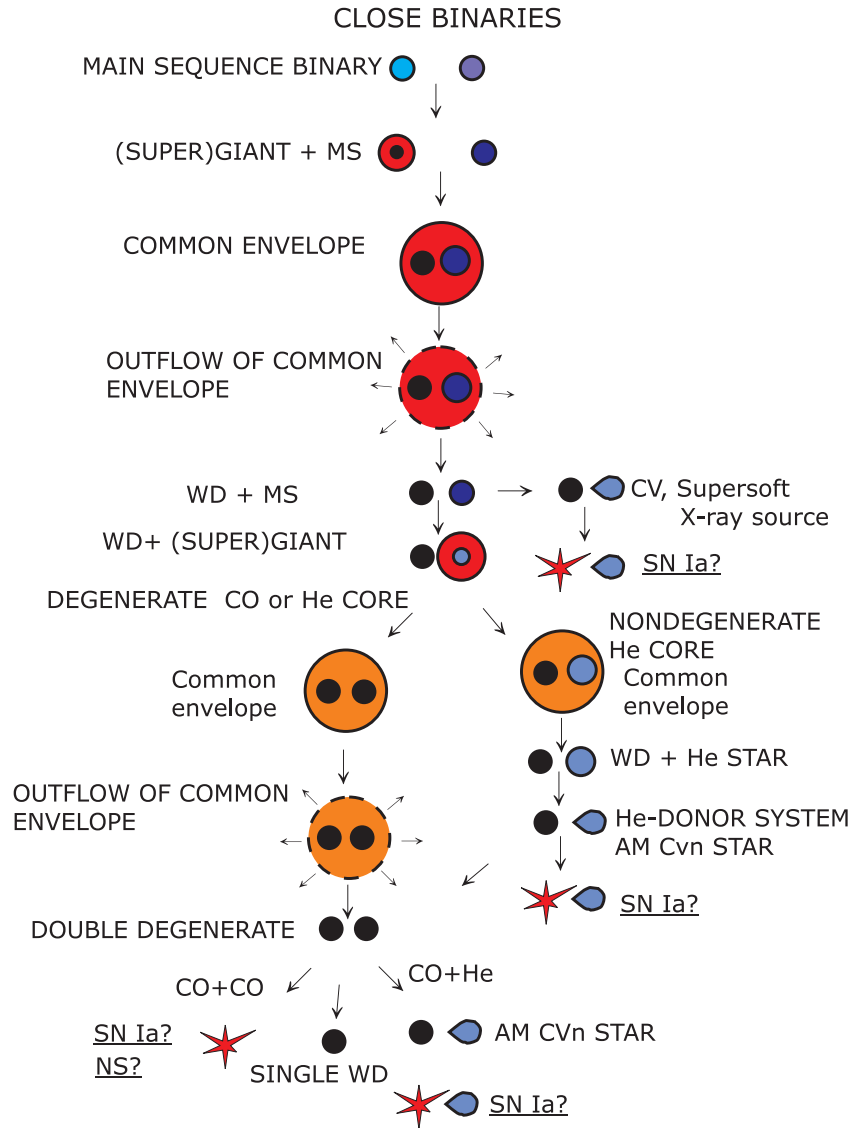


Figure 5: Formation of close binary dwarfs and their descendants (scale and colour-coding are arbitrary).

circularization radius approaches the 2:1 Lindblad resonance. This can efficiently prevent mass being transferred onto the compact object. These endpoints of the evolution of binaries with low-mass donors, were, in fact, never studied.

The second common envelope may form when the companion to the WD overfills its Roche lobe. If the system avoids merger and the donor had a degenerate core, a close binary WD (or double-degenerate, DD) is formed. The fate of the DD is solely defined by GWR. The closest of them may be brought into contact by AML via GWR. The outcome of the contact depends on the chemical composition of the stars and their masses. The lighter of the two stars fills the Roche lobe first (by virtue of the mass–radius relation $R \propto M^{-1/3}$). For a zero-temperature WD the condition of stable mass transfer is $q < 2/3$ (but see the more detailed discussion in Section 9). The merger of the CO-WD pair with a total mass exceeding M_{Ch} may result in a SN Ia leaving no remnant (“double-degenerate SN Ia scenario” first suggested by Webbink [443] and independently by Tutukov and Yungelson [418]) or in an AIC with formation of a single neutron star [259]. The issue of the merger outcome for $M_{\text{tot}} > M_{\text{Ch}}$ still remains an unsolved issue and a topic of fierce discussion, see below. For total masses lower than M_{Ch} the formation of a single WD is expected.

If in a CO + He WD dwarfs pair the conditions for stable mass exchange are fulfilled, an AM CVn system forms (see for details [417, 422, 280, 249, 120]).

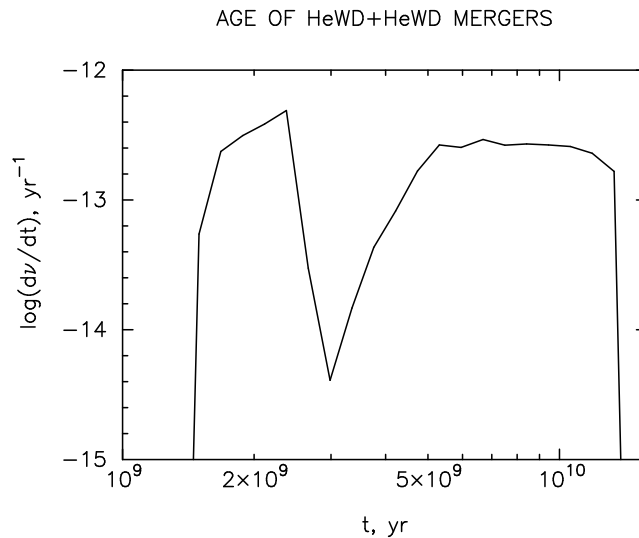


Figure 6: *The age of merging pairs of helium WDs. Two components of the distribution correspond to the systems that experienced in the course of formation two or one common envelope episodes, respectively.*

The current Galactic merger rate of close binary WDs is about 50% of their current birth rate [285, 423]. It is not yet clear how the merger proceeds; it is possible that for He + He or CO + He pairs a helium star is an intermediate stage (see, e.g., [124]). It is important in this respect that binary white dwarfs at birth have a wide range of separations and merger of them may occur gigayears after formation. Formation of helium stars via merger may be at least partially responsible for the ultraviolet flux from the giant elliptical galaxies, where all hot stars finished their evolution long ago. This is illustrated by Figure 6 which shows the occurrence rate of mergers of pairs of He-WDs vs. age.

If the donor has a nondegenerate He-core ($M \gtrsim (2.0-2.5) M_{\odot}$) and the system does not merge, after the second CE-stage a helium subdwarf + WD system may emerge. If the separation of components is sufficiently small, AML via GWR may bring the He-star into contact while He

Donor	CO-WD	MS/SG	He-star	He-WD	RG
Counterpart	Close binary WD	Supersoft XRS	Blue sd	AM CVn	Symbiotic star
Mass transfer mode	Merger	RLOF	RLOF	RLOF	Wind
Young population	10^{-3}	10^{-4}	10^{-4}	10^{-5}	10^{-6}
Old population	10^{-3}	—	—	10^{-5}	10^{-6}

Table 5: Occurrence rates of SNe Ia in the candidate progenitor systems (in yr^{-1}), after [468]. SG stands for sub-giant, RG for red giant, and XRS for X-ray source.

is still burning in its core. If $M_{\text{He}}/M_{\text{wd}} \lesssim 1.2$, stable mass exchange is possible with a typical $\dot{M} \sim 10^{-8} M_{\odot} \text{yr}^{-1}$ [369]. Mass loss quenches nuclear burning and the helium star becomes “semidegenerate”. An AM CVn-type system may be formed in this way (the “nondegenerate He-core” branch of evolution in Figure 5). One cannot exclude that a Chandrasekhar mass may be accumulated by the WD in this channel of evolution, but the probability of such a scenario seems to be very low, $\sim 1\%$ of the inferred Galactic rate of SNe Ia [382]¹¹. If the He-star completes core He-burning before RLOF, it becomes a CO-WD. In Figure 5 it “jumps” into the “double degenerate” branch of evolution.

Type Ia supernovae. Table 5 summarizes order of magnitude model estimates of the occurrence rate of SNe Ia produced via different channels. For comparison, the rate of SNe Ia from wide binaries (symbiotic stars) is also given. The estimates are obtained by a population synthesis code used before in, e.g., [421, 470, 423] for the value of common envelope parameter $\alpha_{\text{ce}} = 1$. The differences in the assumptions with other population synthesis codes or in the assumed parameters of the models result in numbers that vary by a factor of several; this is the reason for giving only order of magnitude estimates. The estimates are shown for $T = 10$ Gyr after the beginning of star formation.

A “young” population had a constant star formation rate for 10 Gyr; in the “old” one the same amount of gas was converted into stars in 1 Gyr. Both populations have a mass comparable to the mass of the Galactic disk. We also list in the table the types of observed systems associated with a certain channel and the mode of mass transfer. These numbers have to be compared to the inferred Galactic occurrence rate of SNe Ia: $(4 \pm 2) \times 10^{-3} \text{yr}^{-1}$ [53]. Table 5 shows that, say, for elliptical galaxies where star formation occurred in a burst, the DD scenario is the only one able to respond to the occurrence of SNe Ia, while in giant disk galaxies with continuing star formation other scenarios may contribute as well.

For about two decades since the prediction of the possibility of the merger of pairs of white dwarfs with total mass $\geq M_{\text{Ch}}$, the apparent absence of observed DDs with proper mass and merger times shorter than Hubble time was considered as the major “observational” difficulty for the DD scenario. Theoretical models predicted that it may be necessary to investigate for binarity up to 1,000 field WDs with $V \lesssim 16-17$ for finding a proper candidate [287]. Currently, it is likely that this problem is resolved (see Section 8).

The merger of pairs of WDs occurs via an intermediate stage in which the lighter of the two dwarfs transforms into a disc [417, 28, 259, 235] from which the matter accretes onto the central object. It was shown for one-dimensional non-rotating models that the central C-ignition and SN Ia explosion are possible only for $\dot{M}_{\text{a}} \lesssim (0.1-0.2)\dot{M}_{\text{Edd}}$ [291, 402]. But it was expected that in the merger products of binary dwarfs \dot{M}_{a} is close to $\dot{M}_{\text{Edd}} \sim 10^{-5} M_{\odot} \text{yr}^{-1}$ [259] because of high viscosity in the transition layer between the core and the disk. For such an \dot{M}_{a} , the nuclear burning will start at the core edge, propagate inward and convert the dwarf into an ONeMg one. The latter will collapse without a SN Ia [167]. However, an analysis of the role of deposition

¹¹The recent discovery of SN Ia SN 2003fg with a mass estimate $\sim 2 M_{\odot}$ [152, 172] may support this scenario.

of angular momentum into a central object by Piersanti and coauthors [315, 316] led them to conclusion that, as a result of the spin-up of rotation of the WD, instabilities associated with rotation, deformation of the WD, and AML by a distorted configuration via GWR, an \dot{M}_a that is initially $\sim 10^{-5} M_\odot \text{ yr}^{-1}$ decreases to $\simeq 4 \times 10^{-7} M_\odot \text{ yr}^{-1}$. For this \dot{M}_a close-to-center ignition of carbon becomes possible. The efficiency of the mechanism suggested in [315, 316] is disputed, for instance, by Saio and Nomoto [364] who found that an off-center carbon ignition occurs even when the effect of stellar rotation is included, if $\dot{M}_a > 3 \times 10^{-6} M_\odot \text{ yr}^{-1}$. The latter authors find that the critical accretion rate for the off-center ignition is hardly changed by the effect of rotation. The problem has to be considered as unsettled until a better understanding of redistribution of angular momentum during the merger process will become available.

Because of a long absence of apparent candidates for the DD scenario and its theoretical problems, the SD scenario is often considered as the most promising one. However, it also encounters severe problems. Even stably burning white dwarfs must have radiatively driven winds. At $\dot{M}_{\text{accr}} \lesssim 10^{-8} M_\odot \text{ yr}^{-1}$ all accumulated mass is lost in nova explosions [331, 461]. Even if \dot{M}_{accr} allows accumulation of a He-layer, most of the latter is lost after the He-flash [163, 54, 314], dynamically or via the frictional interaction of the binary components with the giant-size common envelope. As a result, mass accumulation efficiency is always < 1 and may be even negative. On the other hand, it was noted that the flashes become less violent and more effective accumulation of matter may occur if mass is transferred on a rate close to the thermal one or the dwarf is rapidly rotating [160, 470, 171, 132, 102, 465, 464]. Thus, crucial for this SN Ia scenario are the range of donor masses that may support mass-loss rates “efficient” for the growth of WD, mechanisms for stabilizing mass loss in the necessary range, convection and angular momentum transfer in the accreted layer that define the amount of mass loss in the outbursts, and the amount of matter that escape in the wind¹². If the diversity of SNe Ia is associated with the spread of mass of the exploding objects, it would be more easily explained in a SD scenario, since the latter allows white dwarfs to grow efficiently in mass by shell burning, which is stabilized by accretion-induced spin-up. This inference may be supported by the discovery of the “super-Chandrasekhar” mass SN Ia SN 2003fg (mass estimate $\sim 2 M_\odot$ [152, 172]). Even under assumption of the most favourable conditions for a SN Ia in the SD scenario, the estimates of the current Galactic occurrence rate for this channel do not exceed $1 \times 10^{-3} \text{ yr}^{-1}$ [132], i.e. they may contribute up to 50% of the lowest estimate of the inferred Galactic SN Ia occurrence rate.

On the observational side, the major objection to the SD scenario comes from the fact that no hydrogen is observed in the spectra of SNe Ia, while it is expected that $\sim 0.15 M_\odot$ of H-rich matter may be stripped from the companion by the SN shell [246]¹³. Hydrogen may be discovered both in very early and late optical spectra of SNe and in radio- and X-ray ranges [87, 246, 214]. Panagia et al. [304] find a firm upper limit to a steady mass-loss rate for individual SN systems of $\sim 3 \times 10^{-8} M_\odot \text{ yr}^{-1}$. As well, no expected [246, 52] high luminosity and/or high velocity former companions to exploding WD were discovered as yet¹⁴. The SD scenario also predicts the existence of many more supersoft X-ray sources than are expected from observations, even considering severe problems in estimating incompleteness of the samples of the latter (see for instance [82]).

To summarize, the problem of progenitors of SNe Ia is still unsettled. Large uncertainties in the model parameters involved in the computation of the evolution leading to a SN Ia and in computations of the explosions themselves, do not allow to exclude any type of progenitors. The existence of at least two families of progenitors is suggested by observations (see, e.g., [245]). A high proportion of “peculiar” SN Ia of $(36 \pm 9)\%$ [216] suggests a large spread in the ignition conditions in the exploding objects that also may be attributed to the diversity of progenitors.

¹²All these caveats date back as far as to MacDonalds’ 1984 paper [242] but are still not resolved.

¹³The recently discovered hydrogen-rich SN Ia 2001ic and similarly 1997cy [130] may belong to the so-called SN 1.5 type or occur in a symbiotic system [57].

¹⁴The reported discovery of Tycho Brahe’s 1572 SN Ia companion [360] is not confirmed as yet.

Note that a high proportion of “peculiar” SNe Ia casts a certain doubt to their use as standard candles for cosmology.

As shown in the flowchart in Figure 5, there are configurations for which it is expected that stable accretion of He onto a CO-WD occurs: in AM CVn systems in the double-degenerate formation channel and in precursors of AM CVn systems in the helium-star channel. In the latter systems the mass exchange rate is close to $(1-3) \times 10^{-8} M_{\odot} \text{ yr}^{-1}$, practically irrespective of the combination of donor and accretor mass. It was suggested that in such systems the accumulation of a $\sim 0.1 M_{\odot}$ degenerate He-layer onto a $(0.6-0.9) M_{\odot}$ accretor is possible prior to He-detonation and that the latter initiates a compressional wave that results in the central detonation of carbon [230, 232, 459, 231]; even if central carbon ignition does not occur, the scale of the event is comparable to weak SNe [218, 162, 414]. For a certain time these events involving sub-Chandrasekhar mass accretors (nicknamed “edge-lit detonations”, ELD) that may occur at the rates of $\sim 10^{-3} \text{ yr}^{-1}$ were considered as one of the alternative mechanisms for SNe Ia, although it was shown by Höfflich and Khokhlov [148] that the behaviour of light-curves produced by them does not resemble any of the known SNe Ia. Thus, until recently, the real identification of these events remained a problem. However, it was shown recently by Yoon and Langer [463], who considered angular-momentum accretion effects, that the helium envelope is heated efficiently by friction in the differentially rotating spun-up layers. As a result, helium ignites much earlier and under much less degenerate conditions compared to the corresponding non-rotating case. If the efficiency of energy dissipation is high enough, detonation may be avoided and, instead of a SN, recurrent helium novae may occur. The outburst, typically, happens after accumulation of $0.02 M_{\odot}$. Currently, there is one object known, identified as He-nova – V445 Pup [439, 440, 9, 10]. If He-novae are really associated with mass-transfer from low-mass helium stars to CO white dwarfs, then the estimates of the birth rate of the latter systems ($\approx 0.6 \times 10^{-3} \text{ yr}^{-1}$) [473] and of the amount of matter available for transfer ($\approx 0.2 M_{\odot}$) give an occurrence rate of He-novae of $\sim 0.1 \text{ yr}^{-1}$, i.e. one He-nova per several 100 “ordinary” hydrogen-rich novae.

As we mentioned above, intermediate mass donors, before becoming white dwarfs, pass through the stage of a helium star. If the mass of the latter is above $\simeq 0.8 M_{\odot}$, it expands to giant dimensions after exhaustion of He in the core and may overflow the Roche lobe and, under proper conditions, transfer mass stably [161]. For the range of mass-accretion rates expected for these stars, both the conditions for stable and unstable helium burning may be fulfilled. In the former case the accumulation of M_{Ch} and a SN Ia become possible, as it was shown explicitly by Yoon and Langer [462]. However, the probability of such a SN Ia is only $\sim 10^{-5} \text{ yr}^{-1}$.

Ultra-compact X-ray binaries. The suggested channels for formation of UCXBs in the field are, in fact, “hybrids” of scenarios presented in Figures 4 and 5. In progenitors of these systems, the primary becomes a neutron star, while the secondary is not massive enough. Then, several scenarios similar to the scenarios for the systems with the first-formed white dwarf are open. A white dwarf may overflow the Roche lobe due to systemic AML via GWR. A low-mass companion to a neutron star may overflow the Roche lobe at the end of the main sequence and become a low-mass He-rich donor. A core helium-burning star may be brought in contact by AML due to GWR; mass loss quenches nuclear burning and the donor becomes a helium “semidegenerate” object. An additional scenario is provided by the formation of a neutron-star component by AIC of an accreting white dwarf. We refer the reader to the pioneering papers [411, 288, 394, 369, 412, 101, 434] and to more recent studies [285, 472, 321, 311, 25, 166, 429, 430, 191, 181]. An analysis of the chemical composition of donors in these systems seems to be a promising way for discrimination between systems of different origin [278, 279, 166, 448]: Helium dwarf donors should display products of H-burning, while He-star descendants should display products of He-burning products. Actually, both carbon/oxygen and helium/nitrogen discs in UCXBs are discovered [277].

In globular clusters, UCXBs are formed most probably by dynamical interactions, as first suggested by Fabian et al. [93] (see, e.g., [342, 170, 432, 233, 26] and references therein for the

latest studies on the topic).

8 Observations

The state of interrelations between observations and theoretical interpretations are different for different groups of compact binaries. Such cataclysmic variables as novae stars have been observed for centuries, their lower-amplitude cousins (including AM CVn-type stars) for decades, and their origin and evolution found their theoretical explanation after the role of common envelopes and gravitational waves radiation and magnetic braking were recognized [300, 301, 417, 435]. At present, about 2,000 CVs are known; see the online catalogue by Downes et al. [84] at [383].

More than 600 of them have measured orbital periods; see the online catalogue by Kolb and Ritter [350] at [351].

In particular, there are at present 17 confirmed AM CVn-stars with measured or estimated periods and two more candidate systems; see the lists and references in [272, 390] and [275].

Ultracompact X-ray binaries were discovered with the advent of the X-ray astronomy era in the late 1960s and may be found in the first published catalogues of X-ray sources (see for instance [118]). Their detailed optical investigation became possible only with 8 m-class telescopes. Currently, 12 UCXB systems with measured or suspected periods are known plus six candidates; six of the known systems reside in globular clusters (see lists and references in [277, 20]). Contrary to cataclysmic binaries, the place of UCXBs in the scenarios of evolution of close binaries, their origin and evolution were studied already before optical identification [418, 340, 411].

Unlike CVs and UCXBs, the existence of close detached white dwarfs (DDs) was first deduced from the analysis of scenarios for the evolution of close binaries [443, 417, 418, 160, 444]. Since it was also inferred that DDs may be precursors of SNe Ia, this theoretical prediction stimulated optical surveys for DDs and the first of them was detected in 1988 by Saffer, Liebert, and Olszewski [362]. However, a series of surveys for DDs performed over a decade [352, 41, 108, 248, 363] resulted in only about a dozen of definite detections [247].

8.1 SPY project

The major effort to discover DDs was undertaken by the “ESO Supernovae Ia Progenitors survey (SPY)” project: a systematic radial velocity survey for DDs with the UVES spectrograph at the ESO VLT (with PI R. Napiwotzky; see [265, 266, 372, 268, 267] for the details of the project and [267] for the latest published results). The project was aimed at DDs as potential progenitors of type Ia supernovae, but brought as a by-product an immense wealth of data on white dwarfs. More than 1,000 white dwarfs and pre-white dwarfs were observed (practically all white dwarfs brighter than $V \approx 16.5$ available for observations from the ESO site in Chile). SPY tremendously increased the number of detected DDs to more than 150. Their system parameters are continuously determined from follow-up observations. Figure 7 shows the total masses of the currently known close DDs vs. orbital periods and compares them with the Chandrasekhar mass and the critical periods necessary for merger of components in Hubble time for given M_{tot} (data available in the fall of 2005; R. Napiwotzky, private communication) Altogether, ~ 5 super-Chandrasekhar total mass DDs are expected to be found by SPY. At the moment, several systems with masses close to the Chandrasekhar limit and merger time shorter than Hubble time, including a probable SN Ia progenitor candidate are already detected. The second candidate super-Chandrasekhar mass binary white dwarf was discovered by Tovmassian et al. [404].

Figure 8 shows the position of the observed components of known DDs vs. the theoretical expectations with account for observational selection effects in the $P_{\text{orb}}-m$ diagram [287]. The agreement may be considered as quite satisfactory.

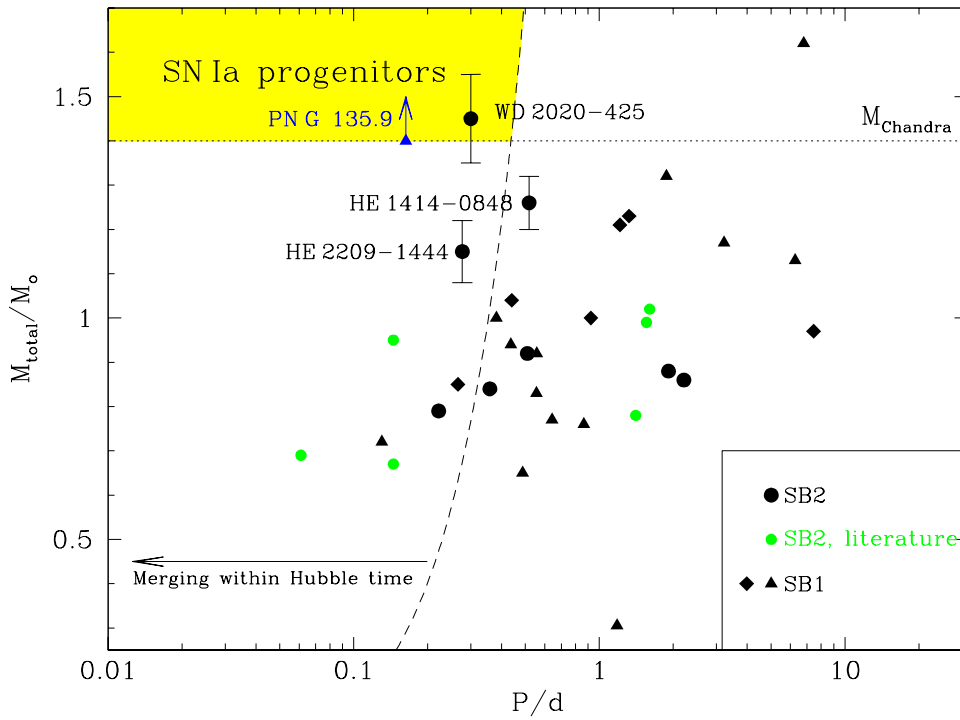


Figure 7: Known close binaries with two WD components, or a WD and a sd component. Green circles mark systems known prior to the SPY project. Black filled symbols mark the positions of DDs and WD + sd systems detected in the SPY project. A blue triangle marks the positions of the WD component of the binary planetary nebula nucleus PN G135.9+55.91 detected by Tovmassian et al. [404]. (Courtesy R. Napiwotzki.)

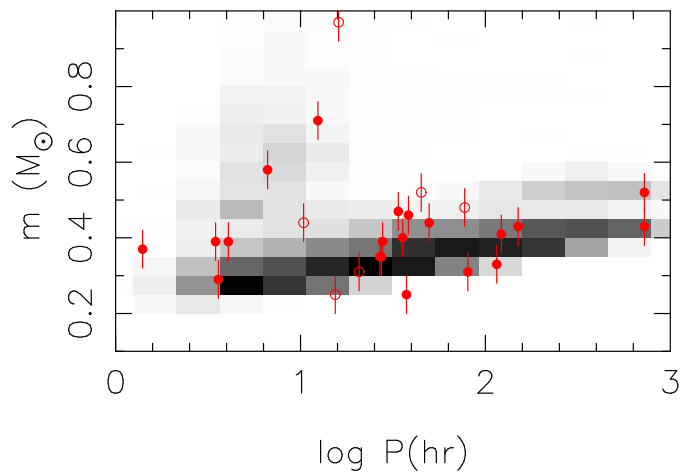


Figure 8: The position of the primary components of known DDs in the “orbital-period–mass” diagram. The underlying gray scale plot is a model prediction from Nelemans et al. [285]. (Figure from [276].)

9 Evolution of Interacting Double-Degenerate Systems

Angular momentum losses via GWR may bring detached double degenerates into contact. The mass–radius relation for degenerate stars has a negative power ($\simeq -1/3$ for WDs with a mass exceeding $0.1 M_{\odot}$, irrespective of their chemical composition and temperature [75]). Hence, the lower mass WD fills its Roche lobe first.

In a binary with stable mass transfer the change of the radius of the donor exactly matches the change of its Roche lobe. This condition combined with an approximation to the size of the Roche lobe valid for low q [194],

$$R_L \approx 0.4622 a \left(\frac{m}{M+m} \right)^{1/3}, \quad (62)$$

where m and M are, respectively, the masses of the prospective donor and accretor, renders two relations important for the study of compact binaries.

1. It provides a relation between the orbital period of the binary P_{orb} and the mass M_2 and radius R_2 of the donor:

$$P_{\text{orb}} \simeq 101 \text{ s} \left(\frac{R_2}{0.01 R_{\odot}} \right)^{3/2} \left(\frac{0.1 M_{\odot}}{M_2} \right)^{1/2}. \quad (63)$$

2. It allows to derive the rate of mass transfer for a semidetached binary in which mass transfer is driven by angular momentum losses:

$$\frac{\dot{m}}{m} = \left(\frac{j}{J} \right)_{\text{GWR}} \times \left(\frac{\zeta(m)}{2} + \frac{5}{6} - \frac{m}{M} \right)^{-1}, \quad (64)$$

where $\zeta(m) = d \ln r / d \ln m$.

For the mass transfer to be stable, the term in the brackets must be positive, i.e.

$$\frac{m}{M} < \frac{5}{6} + \frac{\zeta(m)}{2}. \quad (65)$$

Violation of this criterion results in mass loss by the donor on a dynamical time scale and, most probably, merger of components. Of course, Equation (65) clearly oversimplifies the conditions for stable mass exchange. A rigorous treatment has to include a consideration of tidal effects, angular momentum exchange, and possible super-Eddington \dot{M} immediately after RLOF by the donor and associated common-envelope formation, possible ignition of accreted helium [446, 136, 250, 120]. However, the study of these effects is still in the embryonic state and usually the evolution of IDD or other binaries with WD donors is calculated applying M – R relations, without considering detailed models of WDs (or low-mass helium stars in the appropriate formation channel); see, e.g., [300, 436, 99, 417, 340, 395, 433, 422, 285, 472, 286]. Figure 9 shows examples of the evolution for systems with a helium degenerate donor or a low-mass “semidegenerate” helium star donor and a carbon-oxygen accretor with initial masses that are currently thought to be typical for progenitors of AM CVn systems – 0.2 and $0.6 M_{\odot}$, respectively. For Figure 9 the mass–radius relation for zero-temperature white dwarfs from the article by Verbunt and Rappaport [433] is used; for the low-mass He-star mass–radius relation approximating results of evolutionary computations by Tutukov and Fedorova [409] were used: $R \approx 0.043 m^{-0.062}$ (in solar units). The same equations are applied for obtaining the model of the population of AM CVn-stars discussed in Sections 10, 11, and 12. The mass of the donor in the system may be a discriminator between the formation channels.

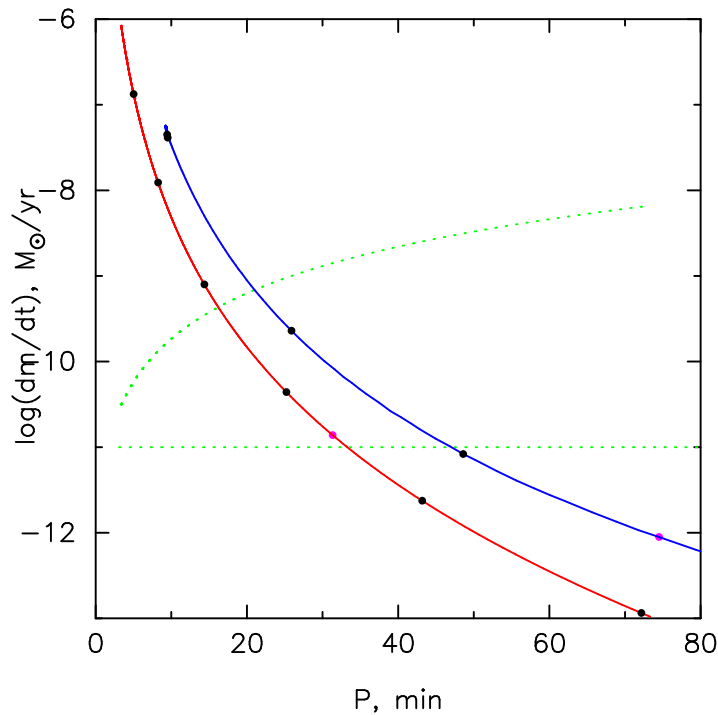


Figure 9: *Mass-loss rate vs. orbital period for “typical” AM CVn-stars: an interacting double degenerate system with initial masses of donor and accretor $0.2 M_{\odot}$ and $0.6 M_{\odot}$ (red line) and a “semidegenerate” low-mass helium star donor plus white dwarf accretor of the same initial masses (blue line). Black dots on the red curve mark positions of the system at $\log T$ (yr) = 5, 6, 7, 8, 9, 10 from Roche-lobe overflow; on the blue curve they mark $\log T$ (yr) = 5, 6, 7, 8. Green lines mark lower and upper limits of the disk instability region according to Tsugawa and Osaki [406]. Below the magenta circles $q < 0.02$ and conventional evolutionary computations may be not adequate for description of mass-transfer process (see the text).*

For instance, a large mass of the donor ($0.18 M_{\odot}$) found for the prototype of the class, AM CVn itself [354] favours the helium channel for this system.

From the Equations (22, 62, 64) it follows that for $m \ll M$ the mass loss rate scales as $M^{1/3}$. As a result, for all combinations of donor and accretor, the $P-\dot{m}$ lines form two rather narrow strips within which they converge with decreasing m . We should note that the time span between formation of a pair of WDs and contact may be from several Myr to several Gyr [422]. This means that the approximation of zero-temperature white dwarfs is not always valid. Below we discuss the implications of finite entropy of the donors for the population of AM CVn-stars.

The “theoretical” model of evolution from shorter periods to longer ones is supported by observations which found that the UV luminosity of AM CVn-stars is increasing as the orbital period gets shorter, since shorter periods are associated with higher \dot{M} [337].

Note that there is a peculiar difference between white dwarf pairs that merge and pairs that start stable mass exchange. The pairs that coalesce stop emitting GWs in a relatively small time-scale (of the order of the period of the last stable orbit, typically a few minutes) [235]. Thus, if we would be lucky to observe a chirping WD and a sudden disappearance of the signal, this will manifest a merger. However, the chance of such event is small since the Galactic occurrence rate of mergers of WDs is $\sim 10^{-2} \text{ yr}^{-1}$ only.

Apart from “double-degenerate” and “helium-star” channels for the formation of AM CVn-stars, there exists the third, “CV”-channel [411, 412, 318]. In this channel, the donor star fills its Roche lobe at the main-sequence stage or just after its completion. For such donors the chemical inhomogeneity inhibits complete mixing at $M \simeq 0.3 M_{\odot}$ typical for initially non-evolved donors. The mixing is delayed to lower masses and as a result the donors become helium dwarfs with some traces of hydrogen. After reaching the minimum period they start to evolve to the longer ones. The minimum of periods for these systems is $\simeq 5-7$ min. However, the birth rate of systems that can penetrate the region occupied by observed AM CVn-stars is much lower than the birth rate in “double degenerate” and “helium-star” channels and we do not take this channel into account below.

10 Gravitational Waves from Compact Binaries with White-Dwarf Components

It was expected initially that contact W UMa binaries will dominate the gravitational wave spectrum at low frequencies [256]. However, it was shown in [417, 252, 160, 92, 222] that it is, most probably, totally dominated by detached and semidetached double white dwarfs.

As soon as it was recognized that the birth rate of Galactic close double white dwarfs may be rather high and even before the first close DD was detected, Evans, Iben, and Smarr in 1987 [92] accomplished an analytical study of the detectability of the signal from the Galactic ensemble of DDs, assuming certain average parameters for DDs. Their main findings may be formulated as follows. Let us assume that there exists a certain distribution of DDs over frequency of the signal f and strain amplitude h : $n(f, h)$. The weakest signal is h_w . For the time span of observations τ_{int} , the elementary resolution bin of the detector is $\Delta f_{\text{int}} \approx 1/\tau_{\text{int}}$. Then, integration of $n(f, h)$ over amplitude down to a certain limiting h and over Δf gives the mean number of sources per unit resolution bin for a volume defined by h . If for a certain h_n

$$\int_{h_n}^{\infty} \left(\frac{dn}{df dh} \right) \Delta \nu_{\text{int}} = 1, \quad (66)$$

then all sources with $h_n > h > h_w$ overlap. If in a certain resolution bin $h_n > h_w$ does not exist, individual sources may be resolved in this bin for a given integration time (if they are above the detector's noise level). In the bins where binaries overlap they produce so-called "confusion noise": an incoherent sum of signals; the frequency, above which the resolution of individual sources becomes possible got the name of "confusion limit", f_c . Evans et al. found $f_c \approx 10$ mHz and 3 mHz for integration times 10^6 s and 10^8 s, respectively.

Independently, the effect of confusion of Galactic binaries was demonstrated by Lipunov, Postnov, and Prokhorov [223] who used simple analytical estimates of the GW confusion background produced by unresolved binaries whose evolution is driven by GWs only; in this approximation, the expected level of the background depends solely on the Galactic merger rate of binary WDs (see [122] for more details). Later, more involved analytic studies of the GW background produced by binary stars at low frequencies were continued in [96, 143, 329, 142].

A more detailed approach to the estimate of the GW background is possible using population synthesis models [222, 445, 285, 286].

Nelemans et al. [285] constructed a model of the gravitational wave signal from the Galactic disk population of binaries containing two compact objects. The model included detached DDs, semidetached DDs, detached systems of NSs and BHs with WD companions, binary NSs and BHs. For the details of the model we refer the reader to the original paper and references therein. Table 6 shows the number of systems with different combinations of components in the Nelemans et al. model¹⁵. Note that these numbers strongly depend on the assumptions in the population synthesis code, especially on the normalization of stellar birth rate, star formation history, distributions of binaries over initial masses of components and their orbital separations, treatment of stellar evolution, common envelope formalism, etc. For binaries with relativistic components (i.e. descending from massive stars) an additional uncertainty is brought in by assumptions on stellar wind mass loss and natal kicks. The factor of uncertainty in the estimated number of systems of a specific type may be up to a factor ~ 10 (cf. [131, 285, 423, 155]). Thus these numbers have to be taken with some caution; we will show the effect of changing some of approximations below. Table 6 immediately shows that detached DDs, as expected, dominate the population of compact binaries.

Population synthesis computations yield the ensemble of Galactic binaries at a given epoch with their specific parameters M_1 , M_2 , and a . Figure 10 shows examples of the relation between

¹⁵These numbers, like also for models discussed below, represent one random realization of the model and are subject to Poisson noise.

Type	Birth rate	Merger rate	Number
Detached DD	2.5×10^{-2}	1.1×10^{-2}	1.1×10^8
Semidetached DD	3.3×10^{-3}	—	4.2×10^7
NS + WD	2.4×10^{-4}	1.4×10^{-4}	2.2×10^6
NS + NS	5.7×10^{-5}	2.4×10^{-5}	7.5×10^5
BH + WD	8.2×10^{-5}	1.9×10^{-6}	1.4×10^6
BH + NS	2.6×10^{-5}	2.9×10^{-6}	4.7×10^5
BH + BH	1.6×10^{-4}	—	2.8×10^6

Table 6: *Current birth rates and merger rates per year for Galactic disk binaries containing two compact objects and their total number in the Galactic disk [285].*

frequency of emitted radiation and amplitude of the signals from a “typical” double degenerate system that evolves into contact and merges, for an initially detached double degenerate system that stably exchanges matter after contact, i.e. an AM CVn-type star and its progenitor, and for an UCXB and its progenitor. For the AM CVn system effective spin-orbital coupling is assumed [280, 249]. For the system with a NS, the mass exchange rate is limited by the Eddington one and excess of the matter is “re-ejected” from the system” (see Section 3.2.3 and [472]). Note that for an AM CVn-type star it takes only ~ 300 Myr after contact to evolve to $\log f = -3$ which explains their accumulation at lower f . For UCXBs this time span is only ~ 20 Myr. In the discussed model, the systems are distributed randomly in the Galactic disk according to

$$\rho(R, z) = \rho_0 e^{-R/H} \operatorname{sech}(z/\beta)^2 \text{ pc}^{-3}, \quad (67)$$

where $H = 2.5$ kpc [361] and $\beta = 200$ pc. The Sun is located at $R_\odot = 8.5$ kpc and $z_\odot = -30$ pc.

Then it is possible to compute strain amplitude for each system. The power spectrum of the signal from the population of binaries as it would be detected by a gravitational wave detector, may be simulated by computation of the distribution of binaries over $\Delta f = 1/T$ wide bins, with T being the total integration time. Figure 11 shows the resulting confusion limited background signal. In Figure 12 the number of systems per bin is plotted. The assumed integration time is $T = 1$ yr. Semidetached double white dwarfs, which are less numerous than their detached cousins and have lower strain amplitude dominate the number of systems per bin in the frequency interval $-3.4 \lesssim \log f$ (Hz) $\lesssim -3.0$ producing a peak there, as explained in the comment to Figure 10.

Figure 11 shows that there are many systems with a signal amplitude much higher than the average in the bins with $f < f_c$, suggesting that even in the frequency range seized by confusion noise some systems may be detectable above the noise.

Population synthesis also shows that the notion of a unique “confusion limit” is an artifact of the assumption of a continuous distribution of systems over their parameters. For a discrete population of sources it appears that for a given integration time there is a range of frequencies where there are both empty resolution bins and bins containing more than one system (see Figure 13). For this “statistical” notion of f_c , Nelemans et al. [285] get the first bin containing exactly one system at $\log f$ (Hz) ≈ -2.84 , while up to $\log f$ (Hz) ≈ -2.3 there are bins containing more than one system.

As we noted above, predictions of the population synthesis models are sensitive to the assumptions of the model; one of the most important is the treatment of common envelopes (see Section 3.5). Figure 14 compares the average gravitational waves background formed by Galactic population of white dwarfs under different assumptions on star formation history, IMF, assumed Hubble time, and treatment of some details of stellar evolution (cf. [285, 286]). The comparison with the work of other authors [141, 445, 374] shows that both the frequency of the confusion limit

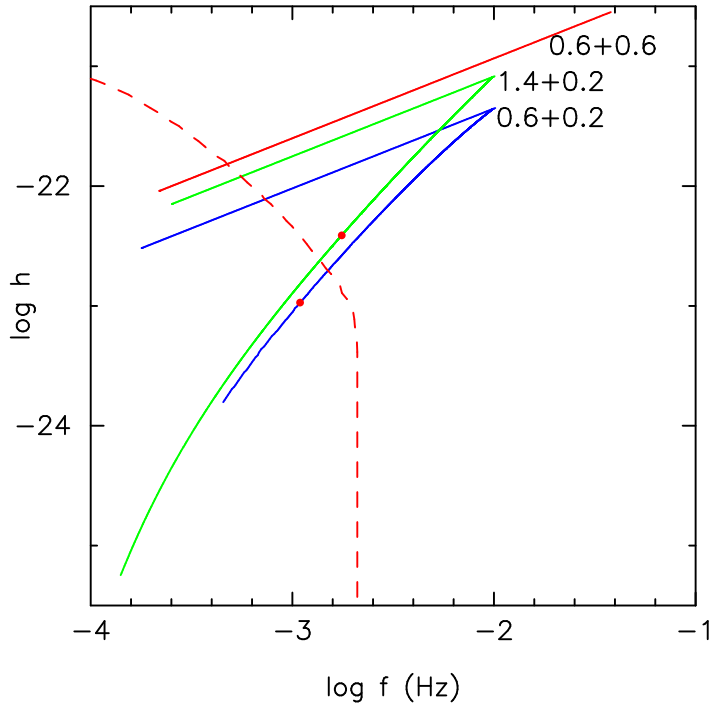


Figure 10: *Dependence of the dimensionless strain amplitude for a WD + WD detached system with initial masses of the components of $0.6 M_{\odot} + 0.6 M_{\odot}$ (red line), a WD + WD system with $0.6 M_{\odot} + 0.2 M_{\odot}$ (blue line) and a WD + NS system with $1.4 M_{\odot} + 0.2 M_{\odot}$ (green line). All systems have an initial separation of components $1 R_{\odot}$ and are assumed to be at a distance of 1 kpc (i.e. the actual strength of the signal has to be scaled with factor $1/d$, with d in kpc). For the DD system the line shows an evolution into contact, while for the two other systems the upper branches show pre-contact evolution and lower branches – a post-contact evolution with mass exchange. The total time-span of evolution covered by the tracks is 13.5 Gyr. Red dots mark the positions of systems with mass-ratio of components $q = 0.02$ below which the conventional picture of evolution with a mass exchange may be not valid. The red dashed line marks the position of the confusion limit as determined in [286].*

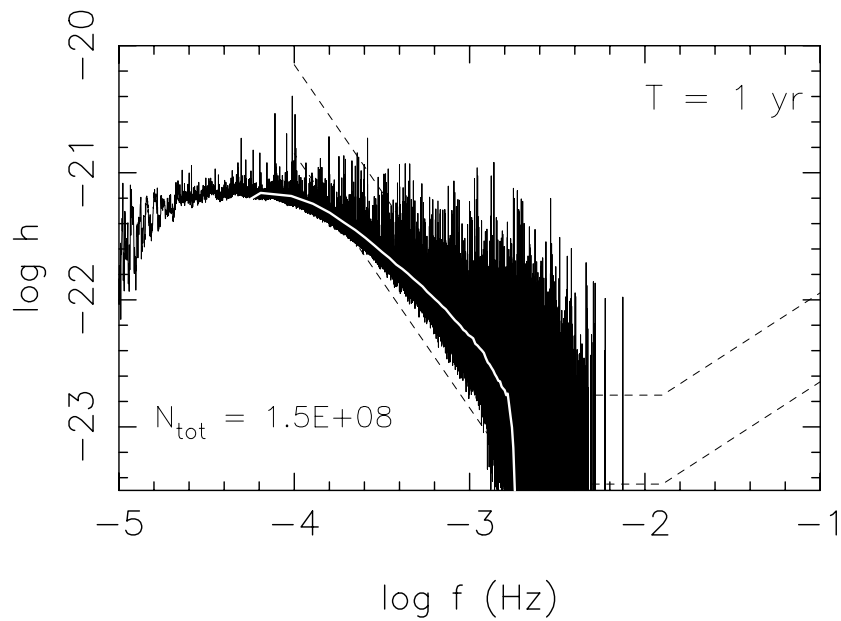


Figure 11: *GWR background produced by detached and semidetached double white dwarfs as it would be detected at the Earth. The assumed integration time is 1 yr. The ‘noisy’ black line gives the total power spectrum, the white line the average. The dashed lines show the expected LISA sensitivity for a S/N of 1 and 5 [210]. Semidetached double white dwarfs contribute to the peak between $\log f \simeq -3.4$ and -3.0 . (Figure from [285].)*

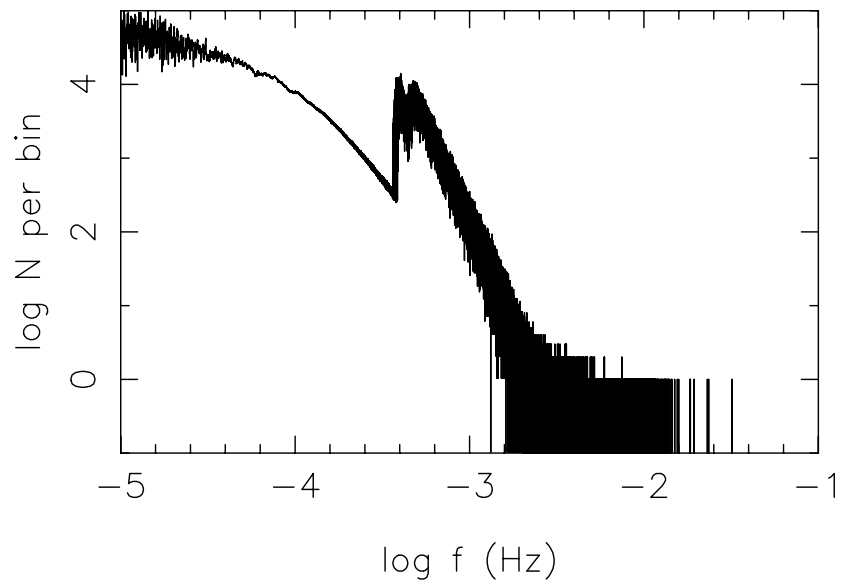


Figure 12: *The number of systems per bin on a logarithmic scale. Semidetached double white dwarfs contribute to the peak between $\log f \simeq -3.4$ and -3.0 . (Figure from [285].)*

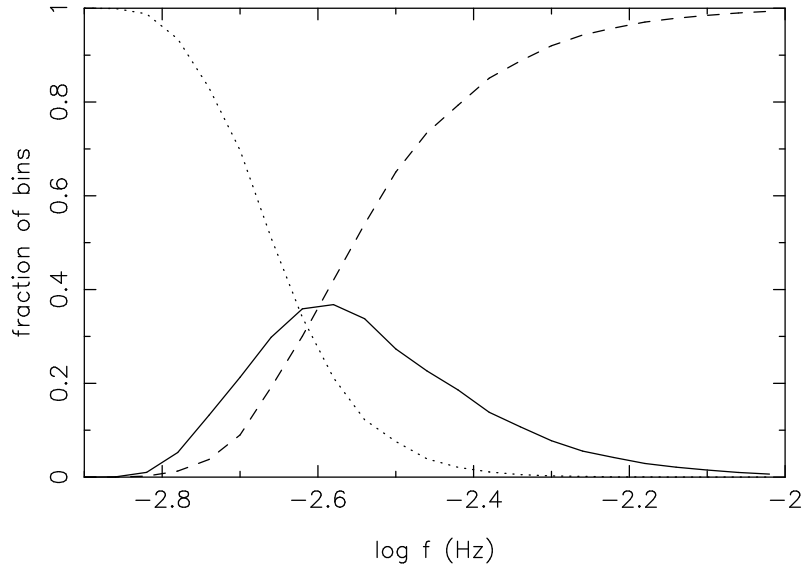


Figure 13: *Fraction of bins that contain exactly one system (solid line), empty bins (dashed line), and bins that contain more than one system (dotted line) as function of the frequency of the signals. (Figure from [285].)*

and the level of confusion noise are uncertain within a factor of ~ 4 . This uncertainty is clearly high enough to influence seriously the estimates of the possibilities for a detection of compact binaries. Note, however, that there are systems expected to be detected above the noise in all models (see below).

Within the model of Nelemans et al. [285] there are about 12,100 detached DD systems that can be resolved above f_c and $\approx 6,100$ systems with $f < f_c$ that are detectable above the noise. This result was confirmed in a follow-up paper [286] which used a more up-to-date SFH and Galaxy model (this resulted in a slight decrease of the number of “detectable” systems – to $\sim 11,000$). The frequency – strain amplitude diagram for DD systems is plotted in the left panel of Figure 15. In the latter study the following was noted. Previous studies of GW emission of the AM CVn systems [142, 285] have found that they hardly contribute to the GW background noise, even despite at $f = (0.3 - 1.0)$ mHz they outnumber the detached DDs. This happens because at these f their chirp mass \mathcal{M} is well below that of a typical detached system. But it was overlooked before that at higher f , where the number of AM CVn systems is much smaller, their \mathcal{M} is similar to that of the detached systems from which they descend. It was shown that, out of the total population of $\sim 140,000$ AM CVn-stars with $P \leq 1500$ s, for $T_{\text{obs}} = 1$ yr, LISA may be expected to resolve $\sim 11,000$ AM CVn systems at $S/N \geq 1$ (or $\sim 3,000$ at $S/N \geq 5$), i.e. the numbers of “resolvable” detached DDs and interacting DDs are similar. Given all uncertainties in the input data, these numbers are in reasonable agreement with the estimates of the numbers of potentially resolved detached DDs obtained by other authors, e.g., 3,000 to 6,000 [374, 62]. These numbers may be compared with expectation that ~ 10 Galactic NS + WD binaries will be detected [61].

The population of potentially resolved AM CVn-type stars is plotted in the right panel of Figure 15. Peculiarly enough, as a comparison of Figures 14 and 15 shows, the AM CVn-type systems appear, in fact, dominant among so-called “verification binaries” for LISA: binaries that are well known from electromagnetic observations and whose radiation is estimated to be sufficiently strong to be detected; see the list of 30 promising candidates in [390] and references therein, and the permanently updated (more rigorous) list of these binaries supported by G. Nelemans, G. Ramsay,

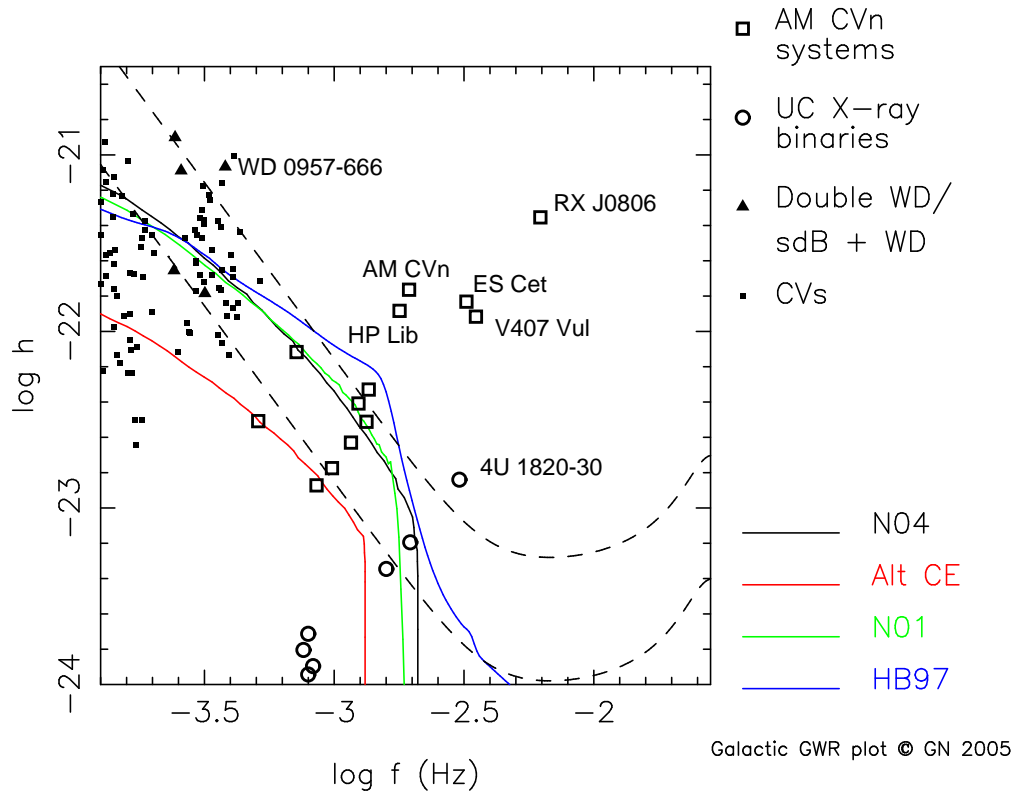


Figure 14: *Gravitational waves background formed by the Galactic population of white dwarfs and signal amplitudes produced by some of the most compact binaries with white dwarf components ([272]). The green line presents results from [285], the black one from [286], while the red one presents a model with all assumptions similar to [286], but with the γ -formalism for the treatment of common envelopes [284, 283] (see also Section 3.5). The blue line shows the background derived in [27]. (Figure from [274].)*

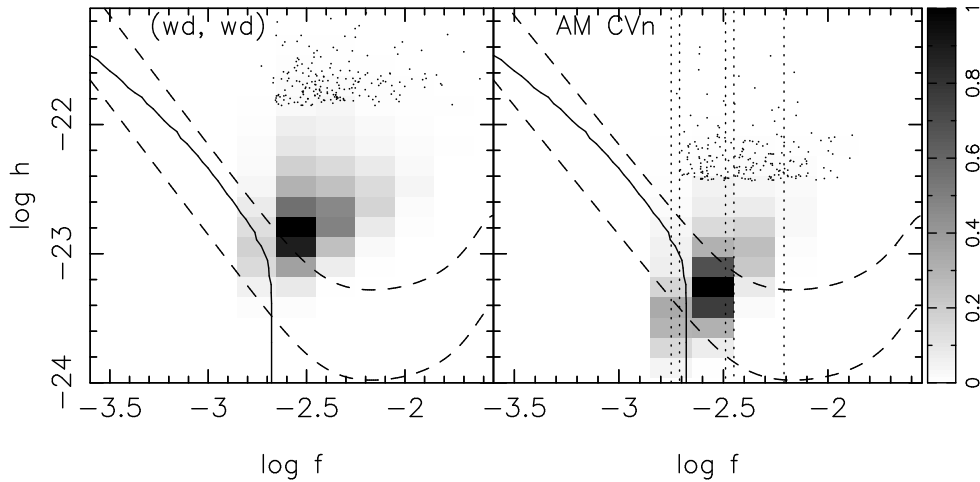


Figure 15: Strain amplitude h as a function of the frequency for the model populations of resolved DDs (left panel, $\simeq 10,700$ objects) and AM CVn systems (right panel, $\simeq 11,000$ objects.). The gray shades give the density distribution of the resolved systems normalized to the maximum density in each panel (1,548 and 1,823 per “cell” for the double white dwarfs and AM CVn-stars panels, respectively). The 200 strongest sources in each sample are shown as dots to enhance their visibility. In the AM CVn panel the periods of several observed short period systems are indicated by the vertical dotted lines. The solid line shows the average background noise due to detached white dwarfs. The LISA sensitivities for an integration time of one year and a signal-to-noise ratio of 1 and 5 are indicated by the dashed lines [210]. (Figure from [286].)

and T. Marsh at [275].

Systems RXJ0806.3+1527, V407 Vul, ES Cet, and AM CVn are currently considered as the best candidates. We must note that the most severe “astronomical” problems concerning “verification binaries” are their distances, which for most systems are only estimates, and poorly constrained component masses.

11 AM CVn-Type Stars as Sources of Optical and X-ray Emission

The circumstances mentioned in the previous paragraph stress the importance of studying AM CVn-stars in all possible wavebands. LISA will measure a combination of all the parameters that determine the GWR signal (frequency, chirp mass, distance, position in the sky, and inclination angle; see, e.g., [139]), so if some of these parameters (period, position) can be obtained from optical or X-ray observations, the other parameters can be determined with higher accuracy. This is particularly interesting for the distances, inclinations, and masses of the systems, which are very difficult to measure with other methods.

In the optical, the total sample of AM CVn-type stars is expected to be dominated by long-period members of the class due to emission of their disks. But the shortest periods AM CVn-type stars that are expected to be observed with LISA may be observed both in optical and X-rays thanks to high mass-transfer rates (see Figure 9). A model for electromagnetic-emission properties of the ensemble of the shortest orbital period $P \leq 1500$ s was constructed by Nelemans et al. [286].

In [286], only systems with He-WD or “semidegenerate” He-star donors were considered (see Figure 5). Systems with donors descending from strongly evolved MS-stars were excluded from consideration, since their fraction in the orbital period range interesting for LISA is negligible. The “optimistic” model of [285] was considered, which assumes efficient spin-orbital coupling in the initial phase of mass-transfer and avoids edge-lit detonations of helium accreted at low \dot{M} . Average temperature and blackbody emission models in V -band and in the ROSAT 0.1–2.4 keV X-ray band were considered, taking into account interstellar absorption. The ROSAT band was chosen because of the discovery of AM CVn itself [424] and two candidate AM CVn systems as ROSAT sources (RXJ0806.3+1527 [168] and V407 Vul [261]) and because of the possibility for a comparison to the ROSAT all-sky survey.

One may identify four main emission sites: the accretion disc and boundary layer between the disc and the accreting white dwarf, the impact spot in the case of direct impact accretion, the accreting star, and the donor star.

Optical emission. The luminosity of the disk may be estimated as

$$L_{\text{disc}} = \frac{1}{2}GM\dot{m} \left(\frac{1}{R} - \frac{1}{R_{L1}} \right) \text{ erg s}^{-1}, \quad (68)$$

with M and R being the mass and radius of the accretor, R_{L1} being the distance of the first Lagrangian point to the centre of the accretor, and \dot{m} being the mass transfer rate, respectively. Optical emission of the disk was modeled as that of a single temperature disc that extends to 70% of the Roche lobe of the accretor and radiates as a blackbody [438].

The emission from the donor was treated as the emission of a cooling white dwarf, using approximations to the cooling models of Hansen [137].

The emission from the accretor was treated as the unperturbed cooling luminosity of the white dwarf¹⁶.

A magnitude-limited sample was considered, with $V_{\text{lim}} = 20$ mag, typical for observed short-period AM CVn-type stars. Interstellar absorption was estimated using Sandage’s model [366] and Equation (67).

X-ray emission. Most AM CVn systems experience a short (10^6 – 10^7 yr) “direct impact” stage in the beginning of mass-transfer [142, 285, 249]. Hence, in modeling the X-ray emission of AM CVn systems one has to distinguish two cases: direct impact and disk accretion.

¹⁶This is true for short-period systems, but becomes an oversimplification at longer periods and for $P_{\text{orb}} \gtrsim 40$ min heating of the WD by accretion has to be taken into account [34].

In the case of a direct impact a small area of the accretor's surface is heated. One may assume that the total accretion luminosity is radiated as a blackbody with a temperature given by

$$\left(\frac{T_{\text{BB}}}{T_{\odot}}\right)^4 = \frac{1}{s} R^{-2} L_{\text{acc}}, \quad (69)$$

where L_{acc} and R are in solar units and L_{acc} is defined by Equation (68). The fraction s of the surface that is radiating depends on the details of the accretion. It was set to 0.001, consistent with expectations for a ballistic stream [238] and the observed X-ray emission of V407 Vul, a known direct-impact system [251].

In the presence of a disk, X-ray emission was assumed to be coming from a boundary layer with temperature [333]

$$T_{\text{BL}} = 5 \times 10^5 \left(\frac{\dot{m}}{10^{18} \text{ g s}^{-1}}\right)^{\frac{2}{9}} \left(\frac{M}{M_{\odot}}\right)^{\frac{1}{3}} \left(\frac{R}{5 \times 10^8 \text{ cm}}\right)^{-\frac{7}{9}} \text{ K}. \quad (70)$$

The systems with an X-ray flux in the ROSAT band higher than $10^{-13} \text{ erg s}^{-1} \text{ cm}^{-2}$ were selected. Then, the intrinsic flux in this band, the distance and an estimate of the Galactic hydrogen absorption [260] provide an estimate of detectable flux.

Figure 16 presents the resulting model. In the top panel there are 220 systems only detectable in X-rays and 330 systems also detectable in the V-band. One may distinguish two subpopulations in the top panel: In the shortest period range there are systems with white-dwarf donors with such high \dot{M} that even sources close to the Galactic centre are detectable. Spatially, these objects are concentrated in a small area on the sky. At longer periods the X-rays get weaker (and softer) so only the systems close to the Earth can be detected. They are more evenly distributed over the sky. Several of these systems are also detectable in the optical (filled symbols). There are 30 systems that are close enough to the Earth that the donor stars can be seen as well as the discs (filled squares). Above $P = 600 \text{ s}$ the systems with helium-star donors show up and have a high enough mass transfer rate to be X-ray sources, the closer ones of which are also visible in the optical, as these systems always have a disc. The bottom panel shows the 1,230 "conventional" AM CVn systems, detectable only by optical emission, which for most systems emanates only from their accretion disc. Of this population 170 objects closest to the Earth also have a visible donor. The majority of the optically detectable systems with orbital periods between 1,000 and 1,500 s are expected to show outbursts due to the viscous-thermal disc instability [406] which could enhance the chance of their discovery.

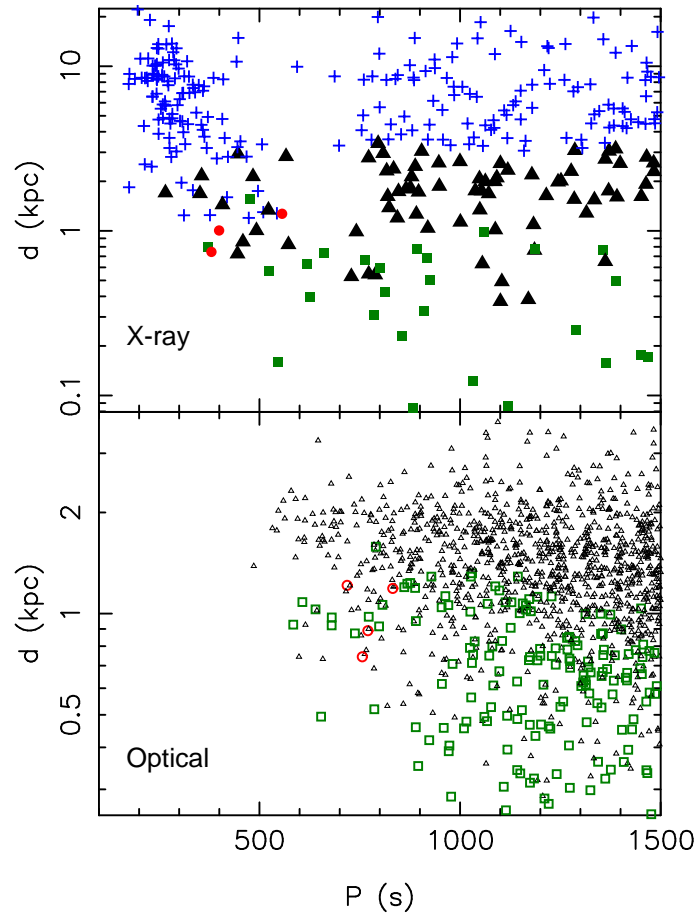


Figure 16: *Distribution of short period AM CVn-type systems detectable in soft X-rays and as optical sources as a function of orbital period and distance. Top panel: systems detectable in X-rays only (blue pluses), direct impact systems observable in X-ray and V-band (red filled circles), systems detectable in X-ray with an optically visible donor (green squares), and systems detectable in X-rays and with an optically visible disc (large filled triangles). Bottom panel: direct impact systems (red open circles), systems with a visible donor (green squares), and systems with a visible accretion disc (small open triangles). The overlap of these systems with systems observable in gravitational waves is shown in Figure 17. (Updated figure from [286], see also [273].)*

12 AM CVn-Type Stars Detectable in GWR and Electromagnetic Spectrum

Figure 17 [286] shows the distributions vs. orbital periods for the total number of AM CVn systems with $P \leq 1500$ s and for AM CVn LISA sources that have optical and/or X-ray counterparts. The interrelations between numbers of sources emitting in different wavebands are shown in the legend to the right of the figure. Out of 11,000 systems detectable in GWR, 2,060 are expected to be in the direct-impact (DI) stage and only 325 in the mass-transfer via disk stage. Thus, the majority of the DI systems are expected to be detectable in GWR; some 5% of DI systems are expected to emit X-rays. There are 1,336 systems detectable in the optical waveband and 326 in X-rays; 106 members of the latter samples may be detected in both spectral ranges.

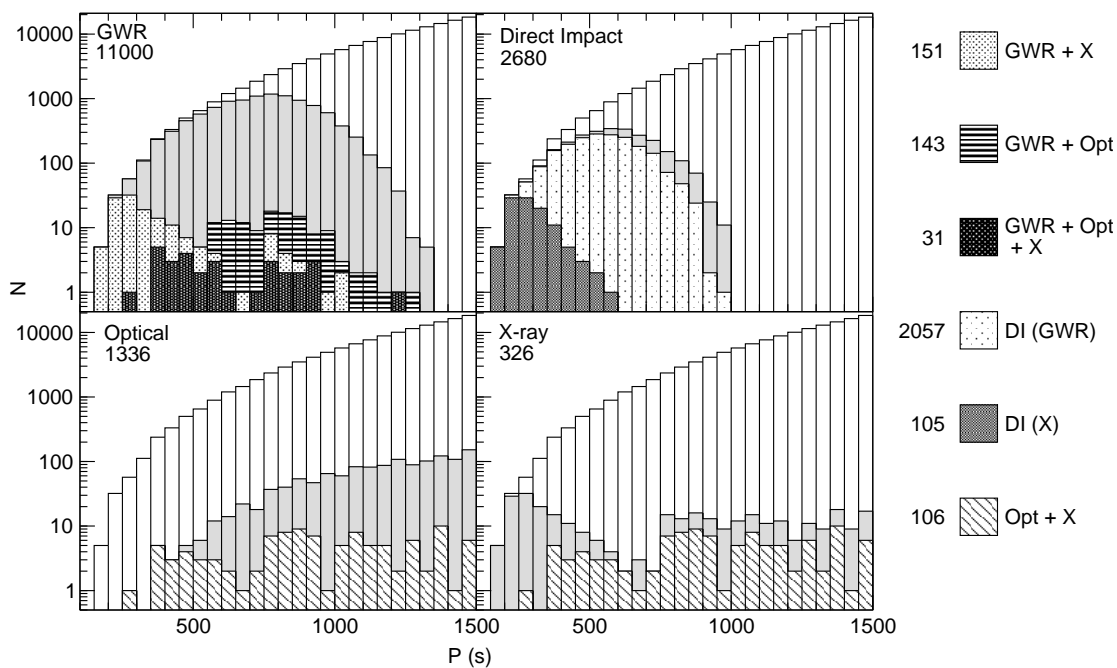


Figure 17: *Short-period AM CVn systems, subdivided in different types. Each panel shows the total population as the white histogram. The top left panel shows 11,000 systems that can be resolved by LISA in gray, and they are subdivided into the ones that have optical counterparts (GWR + Opt), X-ray counterparts (GWR + X), and both (GWR + Opt + X). The top right panel shows the systems that are in the direct impact phase of accretion in gray, and they are subdivided in GWR and X-ray sources. The bottom two panels show (again in gray) the populations that are detectable in the optical band (left panel) and the X-ray band (right panel). The distribution of sources detectable both in optical and X-ray bands is shown as hatched bins in both lower panels (Opt + X). (Figure from [286].)*

An additional piece of information may be obtained from eclipsing AM CVn-stars: They would provide radii of the components and orbital inclinations of the systems. A systematical study of the possibility of eclipses was never carried out, but an estimate for a “typical” system with initial masses of components $(0.25 + 0.60) M_{\odot}$ shows that the probability for eclipsing of the accretor is about 30% at $P = 1000$ s, and even higher for eclipsing (a part of) the accretion disc. The first detection of an eclipsing AM CVn-type star – SDSS J0926+3624 ($P_{\text{orb}} = 28.3$ min) – was recently

reported by Anderson et al. [8].

For WD + WD pairs detectable by LISA the prospects of optical identification are negligible, since for them cooling luminosity is the only source of emission. Most of the potentially detectable dwarfs are located close to the Galactic center and will be very faint. Estimates based on the model [286] predict for the bulk of them $V \approx 35$ mag, with only 75 objects detectable with $V < 25$ mag. Even inclusion of brightening of the dwarfs close to contact under the assumption of efficient tidal heating [164] increases this number to ≈ 130 only (G. Nelemans, private communication).

In the discussion above, we considered the X-ray flux of AM CVn-type systems in the ROSAT waveband: 0.1–2.4 keV. It may be compared with the expected flux in the Chandra and XMM bands: 0.1–15 keV. Since most of the spectra of model AM CVn-stars are rather soft, the flux \mathcal{F} in the latter band is generally not much larger than in the ROSAT band: 80% of systems have $\mathcal{F}_{\text{XMM}}/\mathcal{F}_{\text{ROSAT}} < 1.5$; 96% have $\mathcal{F}_{\text{XMM}}/\mathcal{F}_{\text{ROSAT}} < 3$. However, Chandra and XMM have much higher sensitivity. For instance, the Chandra observations of the Galactic centre have a completeness limit of 3×10^{-15} erg cm $^{-2}$ s $^{-1}$, almost two orders of magnitude deeper than our assumed ROSAT limit [262]. The expected number of X-ray sources in the 0.1–15 keV band detectable down to 10^{-14} erg cm $^{-2}$ s $^{-1}$ is 644 and it is 1085 down to 10^{-15} erg cm $^{-2}$ s $^{-1}$. In the Chandra mosaic image of the Galactic centre [441], roughly down to 10^{-14} erg cm $^{-2}$ s $^{-1}$ there are $\sim 1,000$ point sources, presumably associated with accreting white dwarfs, neutron stars, and black holes. Model [286] predicts 16 X-ray systems in this region.

12.1 Effects of finite entropy

Results concerning the AM CVn-stars population presented above were obtained assuming a mass–radius relation for zero-temperature WDs. Evidently, this is a quite crude approximation, having in mind that in some cases the time span between the emergence of the second white dwarf from the common envelope and contact may be as short as several Myr [422]. As the first step to more realistic models, Deloye and coauthors [76] considered the effects of finite entropy of the donors by using their finite-entropy models for white dwarfs [75]. We illustrate some of these effects following [76].

The effects of finite entropy become noticeably important for $M \lesssim 0.1 M_{\odot}$. Isentropic WD with $T > 0$ (i) have larger radii than $T = 0$ objects and (ii) the M – R relations for them are steeper than for $T = 0$ (i.e. in the range of interest they are still negative but have a lower absolute value). By virtue of Equations (63, 64) this means that for a given orbital period they have higher \dot{M} . This effect is illustrated in the left panel of Figure 18. (The period–mass relation is not single-valued, since the M – R relation has two branches: a branch where the object is thermally supported and a branch where degenerate electrons provide the dominant pressure support.) The right panel of Figure 18 compares a model of the population of AM CVn-stars computed under assumptions that the donor white dwarfs have $T = 0$ and a model which takes into account cooling of the prospective donors between formation and RLOF. The change in the rate of evolution (shown in the left panel) shifts systems with “realistic” cooling to longer orbital periods as compared to the $T = 0$ population.

Finite entropy of the donors also influences the gravitational waves signals from AM CVn-stars. Again, by virtue of the requirement of $R_{\text{donor}} = R_L$, the systems with $T = 0$ donors and hot donors will have a different P_{orb} for the same combination of component masses, i.e. different radii at the contact and different relation between chirp mass \mathcal{M} and P_{orb} . This alters the GW amplitude, Equation (17).

For instance, if a $0.2 M_{\odot}$ donor is fully degenerate, it overflows its Roche lobe at $P_{\text{orb}} \approx 3.5$ min and then evolves to longer P_{orb} . If “realistic” cooling is considered, there are donors that make contact at P_{orb} up to ≈ 25 min. Hotter donors at fixed P_{orb} are more massive, increasing \mathcal{M} and

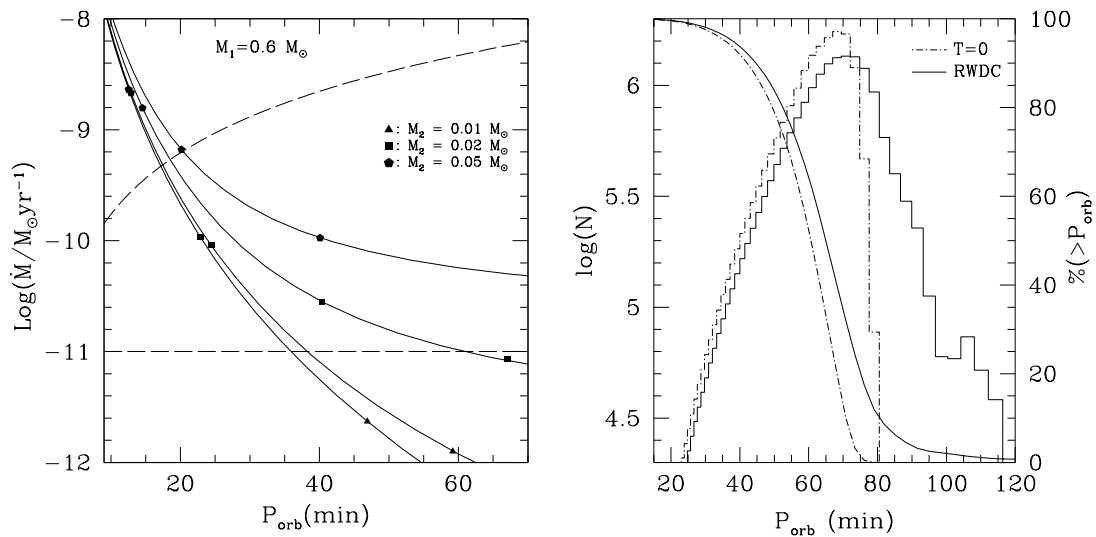


Figure 18: *Effects of a finite entropy of donors on the properties of AM CVn-stars. The left panel shows the relation between P_{orb} and \dot{M} along tracks for a system with initial masses of components $0.2 M_{\odot}$ and $0.6 M_{\odot}$ (like in Figure 9). The solid lines show the evolution for donors with $T_c = 10^4$ K, 10^6 K, 5×10^6 K, and 10^7 K (left to right). The symbols show the positions of models with $M_2 = 0.01 M_{\odot}$ (triangles), $0.02 M_{\odot}$ (squares), and $0.05 M_{\odot}$ (pentagons). The disk stability criteria (for $q = 0.05$) are shown by the dashed lines (after [406]). The right panel compares the numbers of systems as a function of P_{orb} for the model with a $T = 0$ WD (dot-dashed line) and the model with “realistic” cooling (solid lines, RWDC). The smooth curves show the percentage of each population laying above a given P_{orb} . (Figures from [76].)*

increasing h . Thus, the contribution of the individual systems to the integrated GW flux from the total ensemble increases, but their higher rate of evolution decreases the density of population of the sources detectable at low f , since they are lost in the background confusion noise of Galactic WDs. But altogether, the ensemble of sources detectable by LISA with $S/N > 1$ diminishes by about 10% only. Note that finite entropy of donors does not significantly affect the properties of the $\sim 10,000$ systems that are expected to be observed both in electromagnetic spectrum and gravitational waves.

There are some more subtle effects related to finite entropy for which we refer the reader to the original paper.

12.2 Going further

The major issue concerning compact binaries – DDs, IDDs, UCXBs – is their number. Theoretical predictions strongly depend on the assumed parameters and range within an order of magnitude (see references in Section 10). The treatment of common envelopes and the distribution of stars over q are, perhaps, the crux. On the other hand, observational estimates suffer from numerous selection effects. For instance, the estimates of the local space density of white dwarfs may be uncertain by a factor of ~ 5 : cf. $4.2 \times 10^{-3} \text{ pc}^{-3}$ [193], $(5 \pm 0.7) \times 10^{-3} \text{ pc}^{-3}$ [149], or $(20 \pm 7) \times 10^{-3} \text{ pc}^{-3}$ [103]. The problem with AM CVn-stars is their deficiency by a factor of several compared even to the most pessimistic predictions, but in this case the situation improves: about half of the known and candidate systems were found within the past decade, while for another half it took about 30 years. In contrast to DDs, systematic searches for AM CVn-stars started only recently; in fact, the majority of known objects were found serendipitously. Rather successful was the search for AM CVn-candidates in the SDSS catalogue and follow-up observations which brought two confirmed and three candidate stars [353, 8]. Further progress may be expected from dedicated surveys with wide-field cameras. “RApid Temporal Survey” (RATS) by Ramsey and Hakala [338] is aimed at discovering objects with periodicities down to 2 min and is sensitive to $V \sim 22.5$. Preliminary results [336] yielded no new AM CVn-stars, though several objects were expected in the 6 square degrees of the sky covered so far. This may be related to the fact that initially the interstellar extinction for these objects, which have to be located predominantly in the Galactic plane, was underestimated.

Another promising project is “OmegaWhite” (with PI P. Groot, University of Nijmegen in the Netherlands); see [123].

This project is also aimed at compact binaries, it will cover 400 square degrees of the sky, and will provide broad-band and narrow band photometric information down to 21st magnitude or equivalent line fluxes.

Another key issue in the studies of detached and interacting binary white dwarfs is the determination of their distances. In this aspect, the Gaia space probe which is aimed at the creation of a three-dimensional chart of our Galaxy by providing positional and radial velocity measurements for about one billion stars (see [91]) may appear the most promising source of information.

A failure to discover a significant number of detached and interacting double degenerates or to confirm current ideas on their structure and evolution will mean that serious drawbacks exist either in the implementation of known stellar evolution physics and observational statistical data in the population synthesis codes or in our understanding of the processes occurring in compact binaries or in the treatment of selection effects. Special attention in theoretical studies has to be paid to the onset of mass-transfer.

Above, we presented some of the current ideas on formation and compact binaries that may be interesting for general relativity and cosmology and on signals that may be expected from them in the LISA waveband. There is another side of the problem – the analysis of the signal, would it be detected. This topic is out of the scope of this brief review. We refer the reader only to

several papers discussing methods for detecting and subtracting individual binary signals from a data stream with many overlapping signals [63], inferring properties of the distribution of white-dwarf binaries [88], the accuracy of parameter estimation of low-mass binaries [396, 355], and the discussion of the quality of information that may be provided by LISA [391].

13 Conclusions

The current understanding of the evolution of binary stars is firmly based on observations of many types of binary systems, from wide non-interacting pairs to very close compact binaries consisting of remnants of stellar evolution. The largest uncertainties in the specific parameters of the double compact binary formed at the end of the evolution of a massive binary system are related to the physical properties of the pre-supernovae: masses, magnetic fields, equation of state (for NSs), spins, possible kick velocities, etc. This situation is due to our limited understanding of both the late stages of stellar evolution and especially of the supernovae explosion mechanisms. The understanding of the origin and evolution of compact white dwarf binaries also suffers from incompleteness of our knowledge of white dwarf formation and, in particular, on the common envelope treatment. The progress in these fields, both observational and theoretical, will have a major effect on the apprehension of the formation and evolution of compact binary systems. On the other hand, the phenomenological approach used to describe these uncertainties proves to be successful in explaining many observed properties of various binary stars, so the constraints derived from the studies of binary stars should be taken into account in modeling stellar evolution and supernovae explosions.

Of course, specifying and checking the initial distributions of orbital parameters of binary stars and parameters of binary evolution (such as evolution in the common envelopes) stay in the short-list of the important actions to be done. Here an essential role belongs to detailed numerical simulations.

Further observations of compact binaries. Clearly, discoveries of new types of compact binary systems have provided the largest impetus for studies of binary star evolution. The well known examples include the discovery of X-ray binaries, relativistic binary pulsars, and millisecond recycled pulsars. In the nearest future we expect the discovery of NS + BH binaries which are predicted by the massive binary evolution scenario in the form of binary radio pulsars with BH companions [228, 220, 309]. It is very likely that we already observe the coalescence of double NS/BH systems as short gamma-ray bursts in other galaxies [116]. We also anticipate that coalescences of NS + BH or BH + BH binaries can be found first in GW data analysis [122]. The efforts of the LIGO collaboration to put constraints on the compact binary coalescences from the analysis of the existing GW observations are very important [3, 1, 2].

The formation and evolution of compact binaries is a very interdisciplinary field of modern astrophysics, ranging from studies of the equation of state for superdense matter inside neutron stars and testing effects of strong gravity in relativistic compact binaries to hydrodynamical effects in stellar winds and formation of common envelopes. So, further progress in this field, which will be made by means of traditional astronomical observations and new tools, like gravitational wave and neutrino detectors, will undoubtedly have an impact on astronomy and astrophysics as a whole.

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