CHAPTER 1

MASS LOSS AND MASS ACCRETION

Vuoi perdere peso? Chiedimi come![1]
(Herbalife 1995)

1.1. CLOSE BINARY SYSTEMS

In this chapter, one of the fundamental mechanisms for the understanding of a number of phenomena occurring in close binary systems will be described, that is, mass loss from one of the components onto the other.

Also, a review of the way in which the transferred matter sets up and places around the accreting object will be illustrated. Let us however start from the definition of binary system.

One can think that nothing but trivial answers can be given to the question: “When a star is a binary star?”. If one however thinks that a binary system can break up due to the perturbative effect of the closest stars, the question above gains significance: one can thus ask not when a system is binary, but **for how long the two components will form the system**. Chandrasekhar (1944) showed that the che il **break-up time τ_d** of a binary system formed by two stars of mass M_1 and M_2 separated by a distance a is given by the equation

$$\tau_d \cong \frac{(M_1 + M_2)^{\frac{1}{2}}}{4\pi\sqrt{G}mNa^{\frac{3}{2}}} \quad , \tag{1.1}$$

(1) Tr.: *Do you want to lose weight? Ask me how!*

where m is the average mass of the stars close to the system, N their number per unit volume, and G the universal gravitational constant.

Then, following Kopal (1978), one can say that **a system is binary when the probability that the two stars have of completing an orbit is 50%**; alternatively, one can deal with this issue starting from the distance between the two objects. That is, equating the break-up time to the orbital period and using Kepler's third law, one obtains that, in order to consider the two stars actually gravitationally bound, the distance a between them must not overcome the value

$$a_{\max} = 4.81 \cdot 10^4 \left(\frac{M_1 + M_2}{mN} \right)^{\frac{1}{3}} \quad \text{U.A.} \tag{1.2}$$

(for the Solar Neighborhood, $a_{\max}$ is about 130000 A.U., i.e. almost half a parsec). If this value is overcome, the dominant force on the two bodies is not their reciprocal attraction anymore, but instead the sum of the perturbations induced by the nearby stars.

Anyway, using either the first definition of binary system or the second, one can observationally see that binarity is not at all rare in the Galaxy: about one half of the stars which are present in the Galactic Disk are binaries, with peaks of 80% for early-type stars such as O and B (moving toward late-type stars this percentage goes down, even if not strongly); on the contrary, this percentage is much lower (10 - 20%) inside globular clusters, and this because of the high number of stars per cubic parsec which, as already expressed in Eq. (1.1), represents a disgregating factor for the integrity of a binary system.

If the problem is examined from the evolutionary point of view, one can however state that the components of wide binaries can be assimilated to single stars because they pass through their existence independently from each other, without reciprocal influences (that is, without mass exchanges): they could not reach stages in which interactions of great evolutionary importance would take place, not even during the supergiant phase (in the case that one of the two stars, or both, could reach this evolutionary stage), because of the great distance between them.

One can also give a lower limit to the distance itself, as well as to the orbital period, below which a system can be considered a close binary: this limit is around few tens of AU, and corresponds to an orbital period of about one century (these values must however be taken with some degree of caution as they considerably depend on the masses of the two stars and on their mass ratio). Thus, if the system has to be considered close, the history of one of its components cannot be separated by that of the other one. In the next Sections one will briefly see how the mass transfer in close binary systems may take place.

1.2. THE ROCHE MODEL

Generally, when one computes the gravitational interactions between celestial bodies, the approximation of point-like masses, first introduced by Newton, is used. In reality, things are clearly much different, particularly if bodies are at a reciprocal distance which is comparable to the sum of their radii (that is, if they are contact bodies). In this case, the above mentioned approximation cannot be applicable at all, and one must use a method which can describe how the mutual gravitational attraction makes the stars in a close binary system depart from their original spherical shape.

This method, albeit approximate, is called **the Roche model**, named after the French mathematician Edouard Albert Roche who, around the half of XIX century, computed not the equipotential surfaces of two stars independently, but those of a close binary system considered as a whole. The most important characteristic of this model is that these surfaces can be described by means of exact and **purely algebraic** expressions, and therefore relatively easy to be represented.

Three hypotheses are however needed to apply the Roche model:

1) the stars must be **centrally condensed**, that is, most of their mass must be located toward their center and with their central density tending to infinity;
2) orbits must be circular;
3) axial rotation of both stars must be synchronous with the orbital motion and their rotational axes must be perpendicular to the orbital plane.

These hypotheses are essential to maintain the overall fluidodynamic equilibrium in each star (this would not hold anymore if hypothesis [3] fails) and to keep surfaces constant with time (hypothesis [2] is necessary for this reason).
The first hypothesis is instead the one which allows the mathematical description of the Roche model. This description is possible by expressing the problem in spherical polar coordinates using the transformation laws

$$\begin{cases} x = r\cos\varphi\,\mathrm{sen}\,\vartheta = r\lambda \\ y = r\,\mathrm{sen}\,\varphi\,\mathrm{sen}\,\vartheta = r\mu \\ z = r\cos\vartheta \qquad = r\nu \end{cases} \tag{1.3}$$

and is expressed by the formula

$$\xi = \frac{1}{d} + q\left\{\frac{1}{\sqrt{1-2\lambda d}} + \lambda d\right\} - \frac{(q+1)}{2} d^2 \left(1-\nu^2\right) \quad , \tag{1.4}$$

where

$$\xi = \frac{M_2^2}{2M_1(M_1+M_2)} - \frac{R\Psi}{GM_1} \quad , \tag{1.5}$$

$$q = \frac{M_2}{M_1} \quad , \tag{1.6}$$

and

$$d = \frac{r}{R} \tag{1.7}$$

are dimensionless parameters. ξ and Ψ represent the dimensionless potential and the total potential, respectively, acting on the generic point (x, y, z). The parameter q is also called

mass ratio of the system and, as it will be shown further on, it is fundamental in the study of close binaries. For a complete treatment of the Roche model, one can refer to the work by Kopal (1978).

The surfaces over which ξ is constant are called **Roche equipotential surfaces** (Fig. 1.1), and their shape varies according to the value of that quantity.

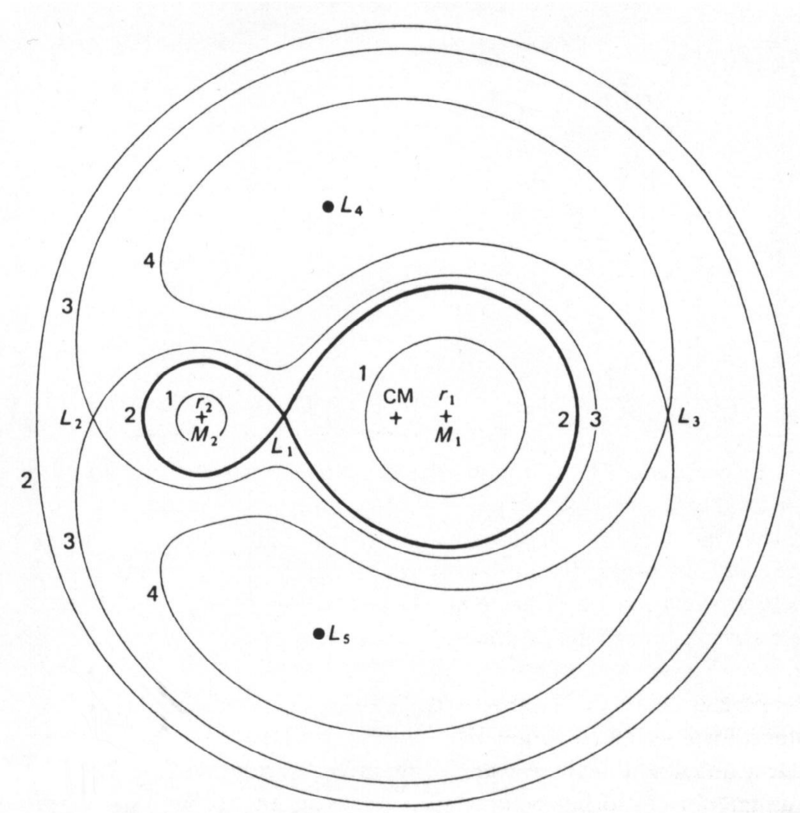

Fig. 1.1. Sections, with respect to the orbital plane, of Roche equipotential surfaces in a binary system with mass ratio q = 0.2; the limit surface is marked with the thick line. The positions of the 5 lagrangian points are also indicated (from: Frank et al. 1992).

If this parameter has an high value, the surface associated to it is almost spherical and centered around the barycenter of star 1, at a small distance from it; on the contrary, as the value of ξ gets smaller, other terms of the expression acquire importance and the surface becomes an oval stretched toward the other star.

The same issue can be applied to the equipotential surfaces of star 2; therefore, as ξ decreases, it will reach the critical value ξ_1 for which the two oval surfaces associated to the two stars will be in contact, thus forming the **Roche limit surface** (Fig. 1.2). This represents the first equipotential surface belonging to both stars.

One can ask why this limit surface is so important. It is because it allows understanding the modalities with which a star which is filling its own **Roche lobe**, that is, the part of limit surface in which the star itself is contained, may lose mass. Indeed, a further expansion of the star would cause, since the surface is equipotential, an overflow of matter from the point in common to the two lobes of the limit surface into the other lobe. This matter will be then almost certainly captured by the other star. So, if one gets back to the answer of the previous Section, it can be stated that **a system is binary when, during its evolution, at least one of the two stars fills its own Roche lobe**.

The contact point between the two lobes is called **inner lagrangian point L_1** and it was called after Giuseppe Luigi Lagrange who, in the second half of the eighteenth century, worked on the algebraic study of the restricted three-body problem (to which the Roche model is in some sense similar; see e.g. Moulton 1970); he found five stationary solutions, the **lagrangian points L_1, L_2, L_3, L_4 and L_5**, placed in well defined positions on the orbital plane (Fig. 1.1).

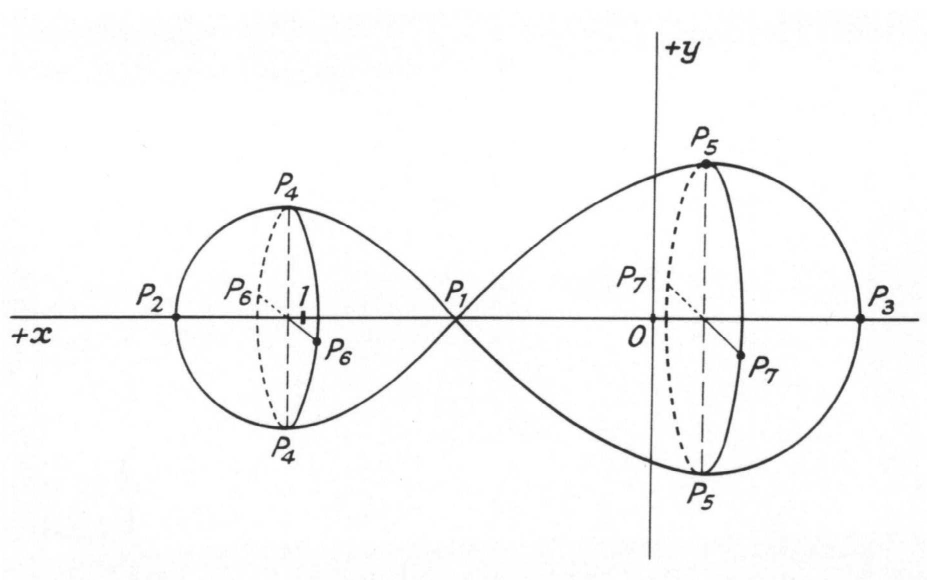

Fig. 1.2. Three-dimensional representation of the Roche limit surface (from: Kopal 1978).

At this time, one can ask how many, in nature, are the systems which satisfy the conditions requested above for this mathematical treatment to be applicable (actually, this question was the reason for which the Roche model has been considered for more than eighty years, until the '40s of this century, a simple theoretical approach to a problem which was considered more tied to fluidodynamics than to astrophysics). Firstly, the central condensation condition has been widely verified during the '20s by Eddington and Emden with their works on the structure of the stellar interiors; then, it has been possible to demonstrate (Kopal 1978) that, in a close binary system, the tidal frictions between the two components tend to circularize their orbits, to synchronize their rotation periods and to to force their rotation axes to be perpendicular to the orbital plane.

There are however cases in which conditions [2] and [3] are not satisfied: therefore, if the orbits are not circular, the shape of the equipotential surfaces will depend on the orbital phase. This would alter the equilibrium condition in the fluid and activate or enhance the mass loss only at particular orbital phases, especially if the eccentricity is high. Instead, if a star has an axial rotation faster than its orbital velocity, it has been calculated that its Roche lobe is smaller than in the case of synchronous rotation. Finally, sometimes it is needed, particularly in X-ray emitting systems, to take into account the pressure exerted by the radiation on the outer layers of the stellar atmospheres: this reduces the gravitational attraction produced by the star by a factor β, which is the ratio between the radiation pressure and the total pressure of the atmosphere itself; also in this case the net result is a shrinking of the star's lobe (Cester 1984).

1.3. MASS LOSS VIA GRAVITATIONAL INSTABILITY

The cases illustrated briefly at the end of the previous Section however need, as a starting point for the interpretation of their scenario (with the exception of the situations in which the hydrostatic equilibrium is strongly violated), of the Roche model and consequently of the coordinates of the points which characterize the limit surface: to this aim, these quantities were tabulated by Plavec & Kratochvil (1964) in dependence on the orbital parameters of the system.

It was noticed that the position of the inner lagrangian point **is a function of the parameter *q* only**, which is the mass ratio between the stars; this favoured the search for approximate formulae for the computation of the position of L_1 with respect to the barycenter of star 1 (i.e. the primary) and of the value R_L of its Roche lobe radius[(2)]. For the latter quantity, Paczynski (1971) found that

$$\frac{R_L}{a} = 0.38 - 0.2 \text{ Log } q \qquad \text{con} \quad 0.3 < \frac{1}{q} < 20 \tag{1.8}$$

and that

$$\frac{R_L}{a} = 0.462 \left(\frac{1}{1+q} \right)^3 \qquad \text{con} \quad 0 < \frac{1}{q} < 0.8 \quad , \tag{1.9}$$

while for the distance b_1 between L_1 and the barycenter of star 1 the following formula holds:

$$\frac{b_1}{a} = 0.5 + 0.227 \text{ Log } q \quad , \tag{1.10}$$

where a is the distance between the two stars. By substituting q with $1/q$, the equivalent quantities for star 2 can be computed. Therefore, a star belonging to a close binary system and expanded during its evolution to a radius similar to or greater than R_L fills its own Roche lobe and **begins to pour matter inside the Roche lobe of its companion**, which in turn begins to accrete this matter. The mass transfer can be quantified by introducing the concept of **mass loss ratio**, defined as

$$\dot{M} = \frac{dM}{dt} \quad , \tag{1.11}$$

which is the lost mass per unit time by the star which fills its Roche lobe. Generally, this quantity is expressed in g s^{-1} or in $M_\odot$ yr^{-1}.

[(2)] The "mean value of the lobe radius" is the value of the radius of the sphere which has the same volume of the considered Roche lobe.

1.4. MASS LOSS VIA STELLAR WIND

The mass transfer through the inner lagrangian point, although it is the more common one, is not the only way in which this phenomenon can take place: there are some other mass loss mechanisms, which however are often of not relevant importance for the problems connected to the intepretation of accretion phenomena in close binary stars. Anyway, in some cases, these alternative mass transfer mechanisms are not negligible at all; the most important and most common of them is the mass loss **via stellar wind**.

Actually, the existence of a stellar wind was first hypothesized to describe a possible reduction of the mass of single stars during their evolution, rather than to explain the effects connected to the accretion in binary systems.

Which are the observational evidences of the stellar wind, and what is the mechanism which causes its origin? Studies performed on the Sun showed evidence of a mass loss via solar wind of about 10^{-14} $M_{\odot}$ yr^{-1} and due to chromospheric activity phenomena. Subsequently these studies were extended to other stars and allowed the discovery that the early-type (O, B and early A) stars are strong wind emitters, as well as the supergiant red stars of the Asymptotic Giant Branch (AGB) of the HR diagram.

Concerning the production mechanism of the stellar wind, the most-credited explanation is the **radiative acceleration mechanism**, which is connected to the action of the radiation pressure P_R, produced by the star radiation on its own outer layers. In the first case mentioned above (that of hot, early-type stars) the ultraviolet emission hits the H and He located in the atmosphere of the star and accelerates them proportionally to T^4, where T indicates the surface temperature of the star: in this way, the outer layer of the star are rapidly accelerated and fastly reach the escape velocity. In supergiant red stars, in a similar manner the infrared radiation (less energetic than the ultraviolet one but very abundant there) acts on the carbon compounds present in the star's atmosphere, which often are condensed in grains due to the low surface temperature. These grains transmit their motion to other atoms and molecules which finally, due to the very low surface gravity acceleration and escape velocity in red supergiants, form an intense wind.

It is believed that also stars of other stellar types, both those still on the Main Sequence and those evolved, emit a particle wind originated by one of these mechanisms (or both, in some cases), but the associated mass loss must be very low as it is not detectable. The stellar wind may also suffer the presence of other stellar phenomena, such as rotation, magnetic field, chromospheric activity or pulsation (of the whole stellar structure or just of the outer layers).

An alternative interpreatation, though less convincing, of the stellar wind genesis is the propagation of acoustic waves in a convective medium, such as the star's envelope.

The mass loss rate for blue giants and red supergiants spans from 10^{-10} $M_\odot$ yr^{-1} to 10^{-6} - 10^{-5} $M_\odot$ yr^{-1}. The accretion via stellar wind is however much less efficient with respect to that via inner lagrangian point because the wind is practically emitted in a isotropic way, and then only a small part (0.1 - 1%) of the expelled mass is captured by the other star; on the contrary, matter which escapes from L_1 falls directly and almost entirely onto the other component of the system. Anyway, it often happens that (particularly in binary systems in which one of the components is an O-type or B-type star; see Sect. 3.3) the mass transfer rate via stellar wind is sufficient to explain the observational data, especially if it is the only allowed one when it is known that the mass losing star does not fill its own Roche lobe.

A wide description of stellar winds may be found in Cassinelli (1979) or, more recently, in Drissen et al. (1992).

1.5. ACCRETION DISKS

Let us now see how accretion takes place after the gas has exited the inner lagrangian point and started its falling on an object of radius R_* and mass M_*. The potential energy released during the falling is, if the mass of the accreted gas is $M_{\rm acc}$,

$$\Delta E_{\text{acc}} = G\frac{M_* M_{\text{acc}}}{R_*} \quad ; \tag{1.12}$$

therefore, if the rate of accreted matter per unit time is $\dot{M}_{\text{acc}}$, the luminosity produced by the falling gas and obtained converting all its potential energy in kinetic energy is

$$L_{\text{acc}} = G\frac{M_* \dot{M}_{\text{acc}}}{R_*} \quad . \tag{1.13}$$

Actually, this description is a first approximation of the problem. It is however appropriate to give a general and precise idea of the luminosities and energies associated to the accretion phenomenon. It is immediately evident that the smaller and the more massive is the accreting object (that is, the higher is its **compactness factor** M_*/R_*) and the higher is the mass accretion rate, the higher will be the luminosity produced with this mechanism.

The emitted spectrum will also depend on these parameters, as well as the gas temperature, which increases because of the internal heating due to viscosity as the gas falls onto the accretor.

The mass accretion via gravitational instability, in which matter passes through the inner lagrangian point and enters the Roche lobe of the accreting star, is the typical example of a non-spherical accretion. In this case, as the orbital velocity of the gas in the lagrangian point L_1 is by far higher (at least of one order of magnitude) than that of free-fall onto the accretor, the net result is that, in close systems, the matter exiting the Roche lobe of the donor star carries a non negligible angular momentum which will keep it from directly falling with an almost radial motion on the accreting star.

The accreted matter will then describe less and less eccentric ellipses around the accretor (this because of the internal gas friction, which will circularize the orbit of the gas particles; indeed, given an orbital momentum $\vec{J}$, the orbit with lower energy is the circular one). Then, an annulus will be formed by the gas at a distance R_{circ} from the attractor, i.e. star 1; an estimate for this distance can be evaluated by supposing that the angular momentum of the gas is constant (at least as a first approximation): one will have, in terms of units of mass for the accreting gas,

$$R_{\mathrm{circ}} \mathrm{v}_{\varphi} = b_1^2 \omega_{\mathrm{orb}} \quad , \tag{1.14}$$

where v_{φ} is the velocity with which matters orbits the accretor at a distance R_{circ}, and ω_{orb} the angular velocity of the binary system. Now, since

$$\mathrm{v}_{\varphi} = \left(\frac{GM_1}{R_{\mathrm{circ}}} \right)^{\frac{1}{2}} \quad , \tag{1.15}$$

and considering $M_* \equiv M_1$, one will have, using once again Kepler's third law and going back to Eq. (1.14), that R_{circ} is ~$0.125a$ (again in the hypothesis that $q \sim 1$), where a is the distance between the centers of the two stars. As one can see, the annulus of gas is well inside the Roche lobe of star 1, given that its lobe has in this case a mean radius of ~$0.38a$, as can be obtained from Eq. (1.8).

If R_{circ} is smaller than the radius R_* of the accreting star, matter falls directly on its surface without completing a single orbit around it: relating R_{circ} (espressed in solar radii), by means of Kepler's third law, with the orbital period of the binary system one can see, again in the hypothesis that $q \sim 1$,

$$R_{\mathrm{circ}} \sim 0.63\, P_{\mathrm{giorni}}^{\frac{2}{3}} \quad R_{\odot} \quad ; \tag{1.16}$$

therefore, if the system is very close and the accretor is a Main Sequence star, the gas falls directly onto its surface without forming particular accretion structures. These structures generally appear if the accreting star is a compact object (white dwarf, neutron star or black hole, as it will be widely shown in the next Chapters). In summary, if the binary system is formed by an ordinary matter star and a collapsed object, the matter lost by the former via gravitational instability through the inner lagrangian point will form an annulus around the compact star.

Due to its own inner viscosity, the gas (the amount of which, meanwhile, increases since other matter is coming from the mass losing star) tends to lose energy and moves to a lower orbit, thus transferring specific angular momentum $J_{\rm sp}$ to the newly-arrived gas from the inner lagrangian point. The new orbit will have a smaller radius R which can be computed from $J_{\rm sp}$ according to the law

$$J_{\rm sp} = \omega_{\rm orb} R^2 = \left(\frac{GM_1}{R^3}\right)^{\frac{1}{2}} R^2 = G^{\frac{1}{2}} M_1^{\frac{1}{2}} R^{\frac{1}{2}} \quad , \tag{1.17}$$

and this phoenomenon rapidly leads to the formation of an **accretion disk** around star 1 (Fig. 1.3), in which the angular momentum lost by the inner parts is transferred to the outer ones which can expand well beyond $R_{\rm circ}$ and, possibly, escape from the system carrying with them a non-negligible part of angular momentum (Pringle 1981); alternatively, they can interact once again with the mass donor star and merge their angular momentum to that associated to the orbital motion of the binary system (Papaloizou & Pringle 1977).

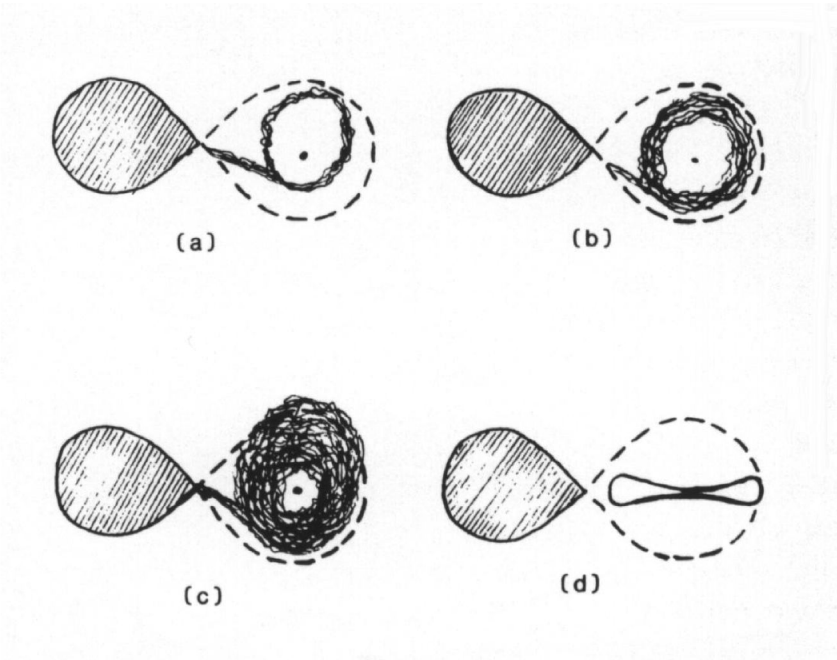

Fig. 1.3. Formation of an accretion disk via mass outflow from the inner lagrangian point. In (d) the section of the system perpendicular to its orbital plane and containing the centers of mass of the two components is represented (from: Petterson 1983).

In addition, the disk is generally **thin**, that is, it has a vertical scale height H for the density that is much smaller than its radius. This is because of the gas internal viscosity and of the tidal friction induced by the accreting star on the gas: these phenomena tend to dissipate the component of the gas motion perpendicular to the orbital plane of the system (which generally is the plane in which the disk forms). If however the accreted matter is not located on the orbital plane, the tidal effects induced by the accretor lead to a precession of the disk and force it to assume a distorted form (Petterson 1983).

All these approximations, and possibly other minor details, allow the description of disk models which are accurate enough to give answers to the problems raised by the observational evidences.

A complete and detailed treatment on the structure and on the properties of accretion disks (the so-called 'standard model') can be found in Frank et al. (1992).

The energy released by the gas in its motion inside the disk is equal to the total energy, with opposite sign, of its last orbit in the disk itself (this because $E_{\mathrm{tot}} \approx 0$, as the matter can be considered to fall from an infinite distance where it has a practically negligible velocity). This means that

$$E_{\mathrm{acc}} = G\frac{M_* m}{R_*} - \frac{1}{2} m \mathrm{v}^2 \quad ; \qquad (1.18)$$

using Eq. (1.15), the previous equation becomes

$$E_{\mathrm{acc}} = \frac{1}{2} G \frac{M_* m}{R_*} \quad , \qquad (1.19)$$

and the luminosity emitted by the disk is therefore

$$L_{\mathrm{disc}} = \frac{1}{2} G \frac{M_* \dot{M}}{R_*} = \frac{1}{2} L_{\mathrm{acc}} \quad . \qquad (1.20)$$

As one can see, half of the accretion luminosity is irradiated by the disk or, more precisely, by the matter which, spiralizing inside it toward the central accreting star, heats up and emits harder and harder radiation. This radiation, if the disk is optically thick, can be absorbed by its outer layers and emitted with a lower frequency. The emission of a large fraction of accretion energy occurs in the place in which the gas stream exiting L_1 hits the outer disk: this place is called **hot spot**. The remaining half of L_{acc} is irradiated in the proximity of the accreting star (Frank et al. 1992).

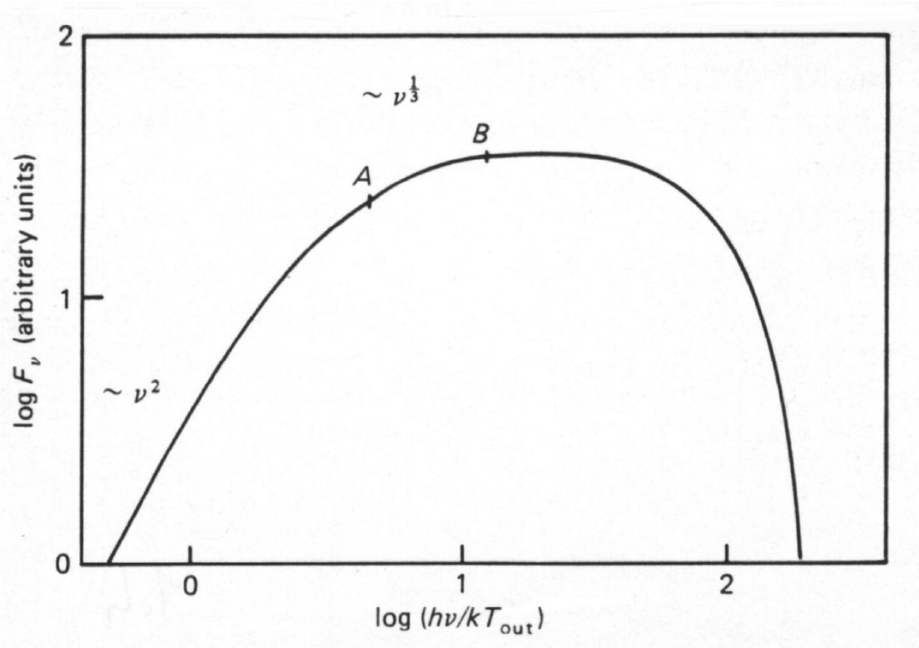

Fig. 1.4. Continuum spectral emission of a stable accretion disk in the hypothesis that it locally irradiates as a blackbody. In the region between points A and B (the optical window) the flux is proportinoal to $\nu^{1/3}$ (from: Frank et al. 1992).

According to the 'standard model', in the hypothesis that the disk locally irradiates as a blackbody, its temperature depends on several quantities and is given by the equation

$$T(R) = \left\{ \frac{3GM\dot{M}}{8\pi R^3 \sigma} \left[1 - \left(\frac{R_*}{R} \right)^{1/2} \right] \right\}^{1/4} \quad . \tag{1.21}$$

At large distances ($R >> R_*$) from the surface of the accreting object, this expression can be simplified as

$$T(R) = T_* \left(\frac{R}{R_*} \right)^{-\frac{3}{4}} \quad , \tag{1.22}$$

where

$$T_* = \left(\frac{3GM\dot{M}}{8\pi R^3 \sigma} \right)^{\frac{1}{4}} \quad ; \tag{1.23}$$

the latter can be considered as a sort of characteristic temperature of the disk.

The spectrum of the disk (Fig. 1.4) generally shows a blue continuum and, in the hypothesis of blackbody emission made before, its flux in the optical region is proportional to $\nu^{1/3}$ (Frank et al. 1992).

1.6. OTHER ACCRETION SHAPES

Let us conclude this overview on the mass loss and the mass accretion with a brief description of the cases in which this latter phenomenon does not take place with a disk shape and the reasons for this different behaviour.

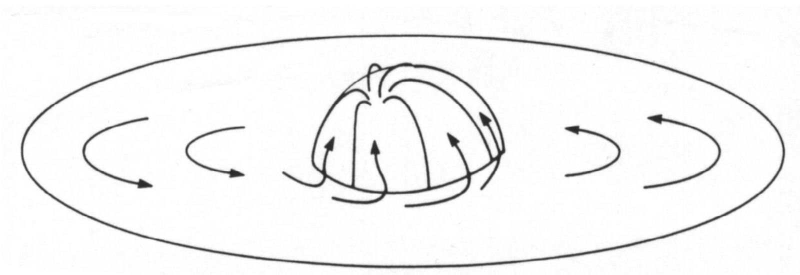

Fig. 1.5. Motion of the gas in an accretion disk around a magnetized compact object (from: Frank et al. 1992).

Often, the motion of the captured matter is influenced not only by the gravitational force, but also by the magnetic field of the accreting object, which can then modify the trajectory of the infalling gas (which is a ionized plasma given its high temperature) leading it to fall onto the magnetic poles of the star following the magnetic field lines. To this aim, it has been defined the **Alfvèn radius** which is, supposing a spherical accretion, the radius of the surface on which the magnetic pressure, proportional to $B^2/8\pi$ (where $\vec{B}$ is the magnetic field), equals the pressure ρv^2 exerted by the falling gas; its value is

$$R_A \cong 2.6 \cdot 10^8 \left(\frac{B_*}{10^{12}\ \text{gauss}}\right)^{\frac{4}{7}} \left(\frac{M_*}{M_\odot}\right)^{\frac{1}{7}} \left(\frac{R_*}{10\ \text{km}}\right)^{\frac{10}{7}} \left(\frac{L_*}{10^{37}\ \text{erg s}^{-1}}\right)^{-\frac{2}{7}} \quad \text{cm} \quad , \qquad (1.24)$$

while, in the disk accretion case, it must be multiplied by a correction factor (Tananbaum & Tucker 1974). Indeed, in compact objects with surface magnetic fields with intensitiy starting from 10^5 - 10^6 gauss the part of the accretion disk which would lie inside the Alfvèn radius is not formed because of the effect induced by $\vec{B}$ (Fig. 1.5). In some cases, if $\vec{B}$ is particularly stong and then R_A is comparable in size with the Roche lobe radius of the compact object, the disk is not formed at all; rather, two **accretion columns** (Fig. 1.6) ending on the magnetic poles of the star are created. This allows the gas hitting only those areas of the compact object surface, creating an highly energetic emission (hard or soft X-rays, depending on the optical depth of the column) along the stellar magnetic axis.

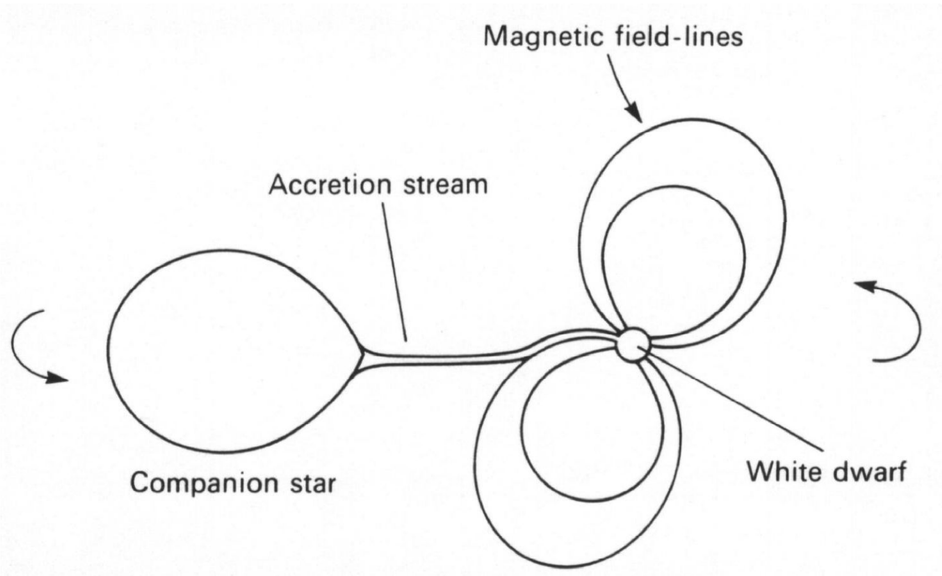

Fig. 1.6. Section of a close binary system showing accretion columns (from: Frank et al. 1992).

Differently, the accretion from stellar wind (Fig. 1.7a,b) produces structures, which will be briefly described here, remarkably different from disks or columns (unless a strong magnetic field is present around the accretor). First of all, one can say that, in the wind case, the star accretes only the gas which has a kinetic energy lower than its potential energy with respect to the accretor itself (Frank et al. 1992); practically, only the matter inside a sphere with radius

$$R_{\mathrm{acc}} \approx \frac{2GM_*}{\mathrm{v}_{\mathrm{rel}}^2} \tag{1.25}$$

(in which

$$\mathrm{v}_{\mathrm{rel}}^2 = \mathrm{v}_{\mathrm{orb}}^2 + \mathrm{v}_{\mathrm{sw}}^2 \tag{1.26}$$

is the square modulus of the star's velocity with respect to the wind, $\vec{\mathrm{v}}_{\mathrm{orb}}$ the orbital velocity vector of the accreting star and $\vec{\mathrm{v}}_{\mathrm{sw}}$ the one referring to the stellar wind), and with center the center of the accreting object, will be accreted. Since very often this relative velocity is supersonic, a shock wave is formed around the object at a distance R_{acc}, from which matter is then captured.

At that distance, the gas has a specific angular momentum, with respect to the accretor, equal to $\omega_{\mathrm{orb}} R_{\mathrm{acc}}{}^2$, therefore much lower that of the matter exiting the inner lagrangian point (Eq. [1.14]), as generally $R_{\mathrm{acc}} << b_1$: thus, in the wind accretion case a disk is hardly formed and accretion takes place with a spherical shape (Fig. 1.7a).

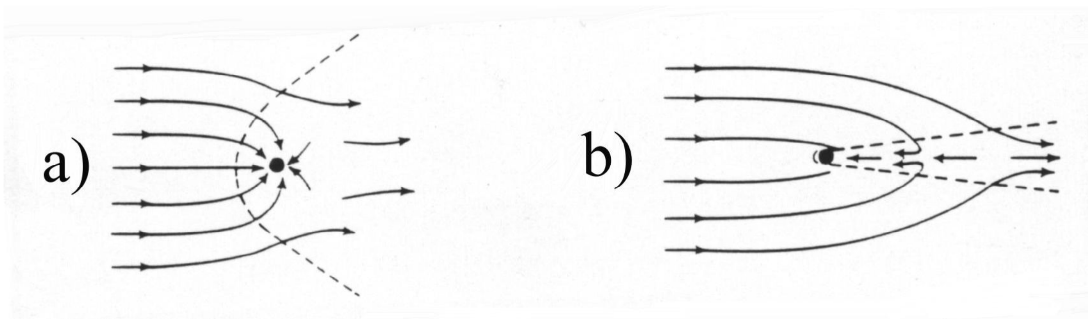

Fig. 1.7. Accretion shapes produced by stellar wind capture: **a** spherical accretion and **b** conical accretion (from: Sunyaev 1978).

If the falling of the gas on the star gives rise to a substantial production of high energy emission, this latter can interact with the gas located on the shock wave, heating it and increasing its velocity: this phenomenon shrinks the accretion sphere, as can be deduced from Eq. (1.25), and creates a **reduced spherical accretion**. Of course, if the accreting object has a strong magnetic field, this may modify the infalling gas motion near the stellar surface and create accretion columns.

Finally, if the stellar wind is particularly dense and slow (30 - 500 km s^{-1}), the shock wave does not form and the gas particles are just deviated by the gravitational field of the accretor: in this case, however, the particles can collide among themselves once they have passed the accreting star, thus dissipating part of their energy and possibly being attracted once again by the star and falling on it if the collision occurs at a sufficiently low distance from the accretor. Here, only the gas contained in a narrow cone with vertex on the star and placed in the opposite direction with respect that of the wind emitter is accreted (**conical accretion**; Fig. 1.7b). For a better comprehension of these phenomena, one can refer to Shakura & Sunyaev (1973, 1975) and Sunyaev (1978).

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