

Hyperons in Neutron Stars

Isaac Vidana

Centro de Física da Universidade de Coimbra (CFisUC), Department of Physics, University of Coimbra, PT-3004-516 Coimbra, Portugal

E-mail: ividana@fis.uc.pt

Abstract. In this work I briefly review some of the effects of hyperons on the properties of neutron and proto-neutron stars. In particular, I revise the problem of the strong softening of the EoS, and the consequent reduction of the maximum mass, induced by the presence of hyperons, a puzzle which has become more intriguing and difficult to solve because of the recent measurements of the unusually high masses of the millisecond pulsars PSR J1903+0327 ($1.667 \pm 0.021 M_{\odot}$), PSR J1614-2230 ($1.97 \pm 0.04 M_{\odot}$), and PSR J0348+0432 ($2.01 \pm 0.04 M_{\odot}$). Some of the solutions proposed to tackle this problem are discussed. Finally, I re-examine also the role of hyperons on the cooling properties of newly born neutron stars and on the so-called r-mode instability.

1. Introduction

Neutron stars are the remnants of the gravitational collapse of massive stars during a Type-II, Ib or Ic supernova event. Their masses and radii are typically of the order of $1 - 2 M_{\odot}$ ($M_{\odot} \simeq 2 \times 10^{33} \text{ g}$ being the mass of the Sun) and $10 - 12 \text{ km}$, respectively. With central densities in the range of $4 - 8$ times the normal nuclear matter saturation density, $\epsilon_0 \sim 2.7 \times 10^{14} \text{ g/cm}^3$ ($\rho_0 \sim 0.16 \text{ fm}^{-3}$), neutron stars are most likely among the densest objects in the Universe [1]. These objects are an excellent observatory to test our present understanding of the theory of strong interacting matter at extreme conditions, and they offer an interesting interplay between nuclear processes and astrophysical observables.

Conditions of matter inside neutron stars are very different from those one can find in the Earth; therefore, a good knowledge of the Equation of State (EoS) of dense matter is required to understand the properties of neutron stars. Nowadays, it is still an open question which is the true nature of neutron stars. Traditionally the core of neutron stars has been modeled as a uniform fluid of neutron-rich nuclear matter in equilibrium with respect to the weak interaction (β -stable matter). Nevertheless, due to the large value of the density, new hadronic degrees of freedom are expected to appear in addition to nucleons. Hyperons, baryons with a strangeness content, are an example of these new degrees of freedom. Contrary to terrestrial conditions, where hyperons are unstable and decay into nucleons through the weak interaction, the equilibrium conditions in neutron stars can make the inverse process happen. Hyperons may appear in the inner core of neutron stars at densities of about $2 - 3\rho_0$. Their presence on the neutron star interior leads to a softening of the EoS and consequently to a reduction of the maximum mass.

Other neutron star properties, such as their thermal and structural evolution, can also be very sensitive to the composition, and therefore to the hyperonic content of neutron star interiors. In particular, the cooling of neutron stars may be affected by the presence of hyperons, since they



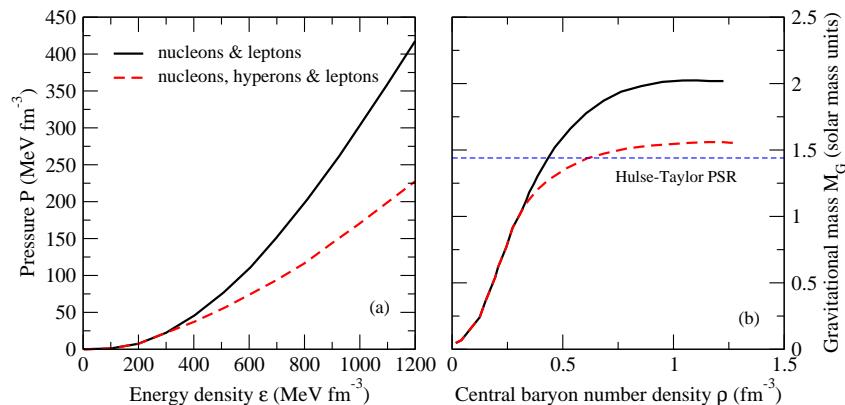


Figure 1. (Color online) Illustration of the effect of the presence of hyperons on the EoS (panel (a)) and mass of a neutron star (panel (b)). A generic model with (black solid line) and without (red dashed line) hyperons has been considered. The horizontal line shows the observational mass of the Hulse–Taylor [14] pulsar.

can modify neutrino emissivities and can allow for fast cooling mechanisms. Furthermore, the emission of gravitational waves in hot and rapidly rotating neutron stars due to the so-called r-mode instability can also be affected by the presence of hyperons in neutron stars, because the bulk viscosity of neutron star matter is dominated by the contribution of hyperons as soon as they appear in the neutron star interior.

In the following I review briefly the so-called hyperon puzzle and other implications of hyperons on the properties of neutron stars.

2. The hyperon puzzle

The presence of hyperons in neutron stars was considered for the first time in the pioneering work of Ambartsumyan and Saakyan in 1960 [2]. Since then, their effects on the properties of these objects have been studied by many authors using either phenomenological [3, 4] or microscopic [5, 6, 7, 8, 9] approaches for the neutron star matter EoS with hyperons. Phenomenological approaches, either relativistic or non-relativistic, are based on effective density-dependent interactions which typically contain a certain number of parameters adjusted to reproduce nuclear and hypernuclear observables, and neutron star properties. Relativistic mean field (RMF) models [3] and Skyrme-type interactions [4] are among the most commonly used ones within this type of approach. Microscopic approaches, on the other hand, are based on realistic two-body baryon-baryon interactions that describe the scattering data in free space. These realistic interactions have been mainly constructed within the framework of a meson-exchange theory [10, 11]. Recently a new approach based on chiral perturbation theory has emerged as a powerful tool [12]. In order to obtain the EoS one has to solve then the many-body problem [13]. However, the repulsive core of the baryon-baryon interaction, which dominates its short-range behavior, is very difficult to treat. Although different microscopic many-body methods have been extensively used to the study of nuclear matter, up to my knowledge, only the Brueckner–Hartree–Fock (BHF) approximation [5] of the Brueckner–Bethe–Goldstone theory, the $V_{low k}$ approach [6], the Dirac–Brueckner–Hartree–Fock theory [7, 8], and very recently the Auxiliary Field Diffusion Monte Carlo method [9], have been extended to the hyperonic sector.

All these approaches agree that hyperons may appear in the inner core of neutron stars at densities of $\sim 2 - 3\rho_0$ as it has been said. At such densities, the nucleon chemical potential is large enough to make the conversion of nucleons into hyperons energetically favorable. This

conversion relieves the Fermi pressure exerted by the baryons and makes the EoS softer, as it is illustrated in panel (a) of Fig. 1 for a generic model with (black solid line) and without (red dashed line) hyperons. As a consequence (see panel (b)) the mass of the star, and in particular the maximum one, is substantially reduced. In microscopic calculations (see *e.g.*, Refs. [5, 6]), the reduction of the maximum mass can be even below the “canonical” one of $1.4 - 1.5M_{\odot}$ [14]. This is not the case, however, of phenomenological calculations for which the maximum mass obtained is still compatible with the canonical value. In fact, most relativistic models including hyperons obtain maximum masses in the range $1.4 - 1.8M_{\odot}$ [3].

Although the presence of hyperons in neutron stars seems to be energetically unavoidable, however, their strong softening of the EoS leads (mainly in microscopic models) to maximum masses not compatible with observation. The solution of this so-called “hyperon puzzle” is not easy, and it is presently a subject of very active research, specially in view of the recent measurements of unusually high masses of the millisecond pulsars PSR J1903+0327 (1.667 ± 0.021) [15], PSR J1614-2230 ($1.97 \pm 0.04M_{\odot}$) [16], and PSR J0348+0432 ($2.01 \pm 0.04M_{\odot}$) [17] which rule out almost all currently proposed EoS with hyperons (both microscopic and phenomenological). To solve this problem it is necessary a mechanism that could eventually provide the additional repulsion needed to make the EoS stiffer and, therefore the maximum mass compatible with the current observational limits. Three different mechanisms that could provide such additional repulsion that have been proposed are: (i) the inclusion of a repulsive hyperon-hyperon interaction through the exchange of vector mesons [18, 19, 20, 21], (ii) the inclusion of repulsive hyperonic three-body forces [22, 23, 24, 25], or (iii) the possibility of a phase transition to deconfined quark matter at densities below the hyperon threshold [27, 28, 29, 30, 31]. In the following I briefly revise these three possible solutions.

2.1. Hyperon-hyperon repulsion

This solution has been mainly explored in the context of RMF models (see *e.g.*, Refs. [18, 19, 20, 21]) and it is based on the well-known fact that, in a meson-exchange model of nuclear forces, vector mesons generate repulsion at short distances. If the interaction of hyperons with vector mesons is repulsive enough then it could provide the required stiffness to explain the current pulsar mass observations. However, hypernuclear data indicates that, at least, the ΛN interaction is attractive [32]. Therefore, in order to be consistent with experimental data of hypernuclei, the repulsion in the hyperonic sector is included only in the hyperon-hyperon interaction through the exchange of the hidden strangeness ϕ vector meson coupled only to the hyperons. In this way, the onset of hyperons is shifted to higher densities and neutron stars with maximum masses larger than $2M_{\odot}$ and a significant hyperon fraction can be successfully obtained. For further information the interested reader is referred to any of the works that have explored this solution in the last years.

2.2. Hyperonic three-body forces

It is well known that the inclusion of three-nucleon forces in the nuclear Hamiltonian is fundamental to reproduce properly the properties of few-nucleon systems as well as the empirical saturation point of symmetric nuclear matter in calculations based on non-relativistic many-body approaches. Therefore, it seems natural to think that three-body forces involving one or more hyperons (*i.e.*, NNY, NYY and YYY) could also play an important role in the determination of the neutron star matter EoS, and contribute to the solution of the hyperon puzzle. These forces could eventually provide, as in the case of the three-nucleon ones, the additional repulsion needed to make the EoS stiffer at high densities and, therefore, make the maximum mass of the star compatible with the recent observations. This idea was suggested even before the observation of neutron stars with $\sim 2M_{\odot}$ (see *e.g.*, Ref. [22]), and it has been explored by some authors in the last years [23, 24, 25]. However, the results of these works show that there is not yet a general

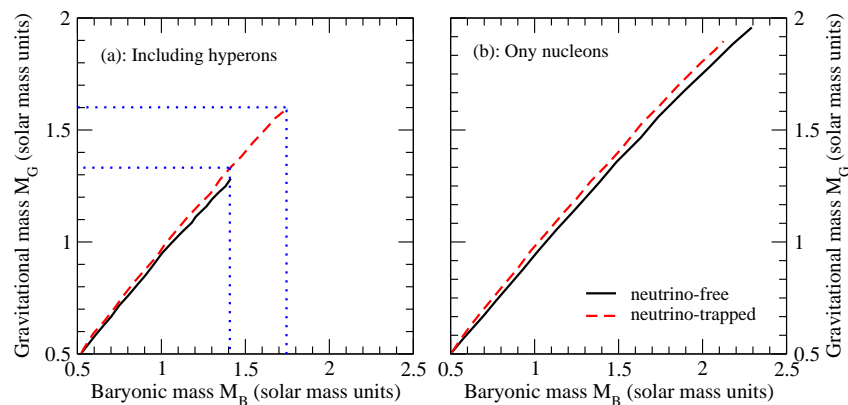


Figure 2. (Color online) Gravitational mass as a function of the baryonic mass for neutrino-free (solid lines) and neutrino-trapped (dashed lines) matter. Panel (a) shows the results for matter containing nucleons and hyperons, whereas the results for pure nucleonic matter are shown in panel (b). Dotted horizontal and vertical lines show the window of metastability in the gravitational and baryonic masses. Figure adapted from Ref. [34].

consensus regarding the role of hyperonic three-body forces on the hyperon puzzle. Whereas in Refs. [22, 24] these forces allow to obtain hyperon stars with $2M_\odot$, in Ref. [23] the larger maximum mass that they can support is $1.6M_\odot$, and the results of Ref. [25] are not conclusive enough because of their strong dependence on the ANN force employed. Therefore, it seems that hyperonic three-body forces are not the full solution to the hyperon puzzle, although, most probably they can contribute to it in a very important way. The interested reader is referred to these works for the specific details of the calculations.

2.3. Quarks in neutron stars

Several authors have suggested that an early phase transition from hadronic matter to deconfined quark matter at densities below the hyperon threshold could provide a solution to the hyperon puzzle. Therefore, massive neutron stars could actually be hybrid stars with a stiff quark matter core. The question that arises in this case is then whether quarks can provide the sufficient repulsion required to produce a $2M_\odot$ neutron star. To yield maximum masses larger than $2M_\odot$, quark matter should have two important and necessary features: (i) a significant overall quark repulsion resulting in a stiff EoS, and (ii) a strong attraction in a particular channel resulting in a strong color superconductivity, needed to make the deconfined quark matter phase energetically favorable over the hadronic one [26]. Several models of hybrid stars with the necessary properties to generate $2M_\odot$ neutron stars have been proposed in the recent years [27, 28, 29, 30, 31]. Conversely, the observation of $2M_\odot$ neutron stars may also helped to impose important constraints on the models of hybrid and strange stars with a quark matter core, and improve our present understanding of the hadron-quark phase transition. Here the interested reader is also referred to the original works for detail information on this possible solution.

3. Hyperon stars at birth and neutron star cooling

Neutron stars are formed after a successful supernova explosion. Properties of newly born neutron stars are affected by thermal effects and neutrino trapping. These two effects have a strong influence on the overall stiffness of the EoS and the composition of the star. In particular (see *e.g.*, [33, 34, 35]) matter becomes more proton rich, the number of muons is significantly

reduced, and the onset of hyperons is shifted to higher densities. In addition, the number of strange particles is on average smaller, and the EoS is stiffer in comparison with the cold and neutrino-free case.

A very important implication of neutrino trapping in dense matter is the possibility of having metastable neutron stars and a delayed formation of a “low-mass” ($M = 1 - 2M_\odot$) black hole. This is illustrated in Fig. 2 for the case of a BHF calculation of Ref. [34]. The figure shows the gravitational mass M_G of the star as a function of its baryonic mass M_B . If hyperons are present (panel (a)), then deleptonization lowers the range of gravitational masses that can be supported by the EoS from about $1.59M_\odot$ to about $1.28M_\odot$ (see dotted horizontal lines in the figure). Since most of the matter accretion on the forming neutron star happens in a very early stages after birth ($t < 1$ s), with a good approximation, the neutron star baryonic mass stays constant during the evolution from the initial proto-neutron star configuration to the final neutrino-free one. Then, for this particular model, proto-neutron stars which at birth have a gravitational mass between $1.28 - 1.59M_\odot$ (a baryonic mass between $1.40 - 1.72M_\odot$) will be stabilized by neutrino trapping effects long enough to carry out nucleosynthesis accompanying a Type-II supernova explosion. After neutrinos leave the star, the EoS is softened and it cannot support anymore the star against its own gravity. The newborn star collapses then to a black hole [33]. On the other hand, if only nucleons are considered to be the relevant baryonic degrees of freedom (panel (b)), no metastability occurs and a black hole is unlikely to be formed during the deleptonization since the gravitational mass increases during this stage which happens at (almost) constant baryonic mass. If a black hole were to form from a star with only nucleons, it is much more likely to form during the post-bounce accretion stage.

The cooling of the newly born hot neutron stars is driven first by the neutrino emission from the interior, and then by the emission of photons at the surface. Neutrino emission processes can be divided into slow and fast processes depending on whether one or two baryons participate. The simplest possible neutrino emission process is the so-called direct Urca process ($n \rightarrow p + l + \bar{\nu}_l$, $p + l \rightarrow n + \nu_l$). This is a fast mechanism which however, due to momentum conservation, is only possible when the proton fraction exceeds a critical value $x_{DURCA} \sim 11\%$ to 15% [36]. Other neutrino processes which lead to medium or slow cooling scenarios but are operative at any density and proton fraction, are the so-called modified Urca processes ($N + n \rightarrow N + p + l + \bar{\nu}_l$, $N + p + l \rightarrow N + n + \nu_l$), the bremsstrahlung ($N + N \rightarrow N + N + \nu + \bar{\nu}$), or the Cooper pair formation ($n + n \rightarrow [nn] + \nu + \bar{\nu}$, $p + p \rightarrow [pp] + \nu + \bar{\nu}$), this last operating only when the temperature of the star drops below the critical temperature for neutron superfluidity or proton superconductivity. If hyperons are present in the neutron star interior new neutrino emission processes, like *e.g.*, $Y \rightarrow B + l + \bar{\nu}_l$, may occur providing additional fast cooling mechanisms. Such additional rapid cooling mechanisms, however, can lead to surface temperatures much lower than that observed, unless they are suppressed by hyperon pairing gaps. Therefore, the study of hyperon superfluidity becomes of particular interest since it could play a key role in the thermal history of neutron stars. Nevertheless, whereas the presence of superfluid neutrons in the inner crust of neutron stars, and superfluid neutrons together with superconducting protons in their quantum fluid interior is well established and has been the subject of many studies, a quantitative estimation of the hyperon pairing has not received so much attention, and just few calculations exists in the literature [37].

4. Hyperons and the r-mode instability of neutron stars

It is well known that the upper limit on the rotational frequency of a neutron star is set by its Kepler frequency Ω_{Kepler} , above which matter is ejected from the star’s equator [38]. However, a neutron star may be unstable against some perturbations which prevent it from reaching rotational frequencies as high as Ω_{Kepler} , setting, therefore, a more stringent limit on its rotation [39]. Many different instabilities can operate in a neutron star. Among them, the so called r-

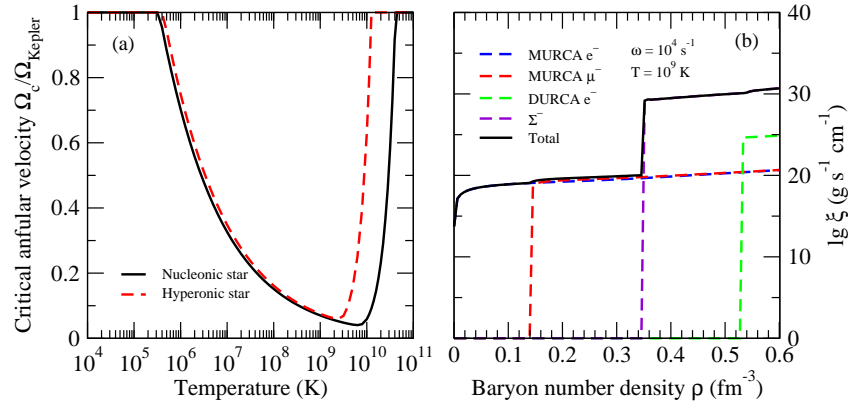


Figure 3. (Color online) Panel (a): r-mode instability region for a pure nucleonic and a hyperonic star with $1.27M_{\odot}$. The frequency of the mode is taken as $\omega = 10^4 \text{ s}^{-1}$. Panel (b): Bulk viscosity as a function of the density for $T = 10^9 \text{ K}$ and $\omega = 10^4 \text{ s}^{-1}$. Contributions direct and modified nucleonic Urca processes as well as from the weak non-leptonic process $n + n \leftrightarrow p + \Sigma^-$ are included. Figure adapted from Ref. [43].

mode instability [40], a toroidal mode of oscillation whose restoring force is the Coriolis force, is particularly interesting. This oscillation mode leads to the emission of gravitational waves in hot and rapidly rotating neutron stars through the Chandrasekhar–Friedman–Schutz mechanism [41]. Gravitational radiation makes an r-mode grow, whereas viscosity stabilizes it. Therefore, an r-mode is unstable if the gravitational radiation driving time is shorter than the damping time due to viscous processes. In this case, a rapidly rotating neutron star could transfer a significant fraction of its rotational energy and angular momentum to the emitted gravitational waves. These waves, potentially detectable, could provide invaluable information on the internal structure of the star and constraints on the EoS.

Bulk (ξ) and shear (η) viscosities are usually considered the main dissipation mechanism of r- and other pulsation modes in neutron stars. Bulk viscosity is the dominant one at high temperatures ($T > 10^9 \text{ K}$) and, therefore, it is important for hot young neutron stars. It is produced when the pulsation modes induce variations in pressure and density that drive the star away from β -equilibrium. As a result, energy is dissipated as the weak interaction tries to reestablish the equilibrium. In the absence of hyperons or other exotic components, the bulk viscosity of neutron star matter is mainly determined by the reactions of direct and modified Urca processes. However, as soon as hyperons appear new mechanisms such as weak non-leptonic hyperon reactions ($N + N \leftrightarrow N + Y$, $N + Y \leftrightarrow Y + Y$), direct ($Y \rightarrow B + l + \bar{\nu}_l$, $B + l \rightarrow Y + \nu_l$) and modified hyperonic Urca ($B' + Y \rightarrow B' + B + l + \bar{\nu}_l$, $B' + B + l \rightarrow B' + Y + \nu_l$), or strong interactions ($Y + Y \leftrightarrow N + Y$, $N + \Xi \leftrightarrow Y + Y$, $Y + Y \leftrightarrow Y + Y$) contribute to the bulk viscosity and dominate it for $\rho > 2 - 3\rho_0$. Several works have been devoted to the study of the hyperon bulk viscosity [42]. The interested reader is referred to these works for detailed studies.

The time dependence of an r-mode oscillation is given by $e^{i\omega t - t/\tau}$, where ω is the frequency of the mode, and τ is an overall time scale of the mode which describes both its exponential growth due to gravitational wave emission as well as its decay due to viscous damping. It can be written as $1/\tau(\Omega, T) = -1/\tau_{\text{GW}} + 1/\tau_{\xi} + 1/\tau_{\eta}$. If τ_{GW} is shorter than both τ_{ξ} and τ_{η} the mode will exponentially grow, whereas in the opposite case it will be quickly damped away. For each star at a given temperature T one can define a critical angular velocity Ω_c as the smallest root of the equation $1/\tau(\Omega_c, T) = 0$. This equation defines the boundary of the so-called r-mode instability region. A star will be stable against the r-mode instability if its angular velocity is smaller than

its corresponding Ω_c . On the contrary, a star with $\Omega > \Omega_c$ will develop an instability that will cause a rapid loss of angular momentum through gravitational radiation until its angular velocity falls below the critical value. On panel (a) of Fig. 3 it is presented, as example, the r-mode instability region for a pure nucleonic (black solid line) and a hyperonic (red dashed line) star with $1.27M_\odot$ [43]. The contributions to the bulk viscosity from direct and modified nucleonic Urca processes as well as from the weak non-leptonic process $n + n \leftrightarrow p + \Sigma^-$ included in the calculation are shown in the panel (b) of the figure. Clearly the r-mode instability is smaller for the hyperonic star. The reason being simply the increase of the bulk viscosity due to the presence of hyperons which makes the damping of the mode more efficient.

5. Summary

I have briefly reviewed the main effects of hyperons on the properties of neutron. In particular, I have revised the problem of the strong softening of the EoS, and the consequent reduction of the maximum mass, due to the presence of hyperons. I have discussed three different solutions proposed to tackle this problem: (i) the inclusion of a repulsive hyperon-hyperon interaction through the exchange of vector mesons (ii) the inclusion of repulsive hyperonic three-body forces, and (iii) the possibility of a phase transition to deconfined quark matter at densities below the hyperon threshold. The role of hyperons on the properties of proto-neutron stars, as well as on the cooling properties of newly born neutron stars, and on the so-called r-mode instability has been also re-examined.

Acknowledgments

The author is very grateful to the organizers of SQM 2015 for their invitation and to D. Unkel for interesting discussions. This work is partly supported by the project PEst-OE/FIS/UI0405/2014 developed under the initiative QREN financed by the UE/FEDER through the program COMPETE-Programa Operacional Factores de Competitividade, and by NewCompstar, COST Action MP1304.

References

- [1] Shapiro S L and Teukolsky S A, *Black Holes, White Dwarf and Neutron Stars* (Wiley, New York) 1983; Glendenning N K, *Compact Stars: Nuclear Physics, Particle Physics and General Relativity*, 2nd edition (Springer, Berlin) 2000; Haensel P, Potekhin A Y and Yakovlev D G, *Neutron Stars 1: Equation of State and Structure, Astrophysics and Space Science Library Series*, Vol. **326** (Springer, New York) 2007.
- [2] Ambartsumyan V A and Saakyan G S 1960 Sov. Astron. **4** 187.
- [3] Glendenning N K 1982 Phys. Lett. B **114**, 392; Glendenning N K 1985 Astrophys. J. **293**, 470; Glendenning N K 1987 Z. Phys. A **326**, 57; Glendenning N K and Moszkowski S A 1991 Phys. Rev. Lett. **67**, 2414; Weber F and Weigel M K 1989 Nucl. Phys. A **505**, 779. Knorren R, Prakash M and Ellis P J 1995 Phys. Rev. C **52**, 3470; Schaffner J and Mishustin I 1996 Phys. Rev. C **53**, 1416; Huber H, Weber F, Weigel M K and Schaab Ch 1998 Int. J. Mod. Phys. E **7**, 310.
- [4] Balberg S and Gal A 1997, Nucl. Phys. A **625**, 435; Balberg S, Lichtenstadt I and Cook G B 1999 Astrophys. J. Suppl. Ser. **121**, 515; Lansky D E and Yamamoto Y 1997 Phys. Rev. C **55**, 2330; Tretyakova T Y and Lansky D E 1999 Eur. Phys. J. A **5**, 391; Cugnon J, Lejeune A and Schulze H-J 2000 Phys. Rev. C **62**, 064308; Vidaña I, Polls A, Ramos A and Schulze H-J 2001 Phys. Rev. C **64**, 044301; Zhou X-R, Schulze H-J, Sagawa H, Wu C-X and Zhao E-G 2007 Phys. Rev. C **76**, 034312; Zhou X-R, Polls A, Schulze H-J and Vidaña I 2008, Phys. Rev. C **78**, 054306.
- [5] Schulze H-J, Baldo M, Lombardo U, Cugnon J and Lejeune A 1995 Phys. Lett. B **355** 21; Schulze H-J, Baldo M, Lombardo U, Cugnon J and Lejeune A 1998 Phys. Rev. C **57** 704; Baldo M, Burgio G F and Schulze H-J 1998 Phys. Rev. C **58** 3688; Baldo M, Burgio G F and Schulze H-J 2000 Phys. Rev. C **61** 055801; Vidaña I, Polls A, Ramos A, Hjorth-Jensen M and Stoks V G J 2000 Phys. Rev. C **61** 025802; Vidaña I, Polls A, Ramos A, Engvik L and Hjorth-Jensen M 2000 Phys. Rev. C **62** 035801; Schulze H-J, Polls A, Ramos A and Vidaña I 2006 Phys. Rev. C **73** 058801; Schulze H-J and Rijken T 2011 Phys. Rev. C **84** 035801.
- [6] Djapo H, Schaefer B-J and Wambach J 2010 Phys. Rev. C **81** 035803.

- [7] Sammarruca F 2009 Phys. Rev. C **79** 034301.
- [8] Katayama T and Saito K 2014 arXiv:1410.7166; 2015 arXiv:1501.05419.
- [9] Lonardoni D, Pederiva F and Gandolfi S 2014 Phys. Rev. C **89** 014314.
- [10] Maessen P M M , Rijken T A and de Swart J J 1989 Phys. Rev. C **40**, 2226; Rijken Th A, Stoks V G J and Yamamoto Y 1999 Phys. Rev. C **59**, 21; Stoks V G J and Rijken Th A 1999 Phys. Rev. C **59**, 3009; Rijken Th A 2006 Phys. Rev. C **73**, 044007; Rijken Th A and Yamamoto Y 2006 Phys. Rev. C **73**, 044008.
- [11] Holzenkamp B, Holinde K and Speth J 1989 Nucl. Phys. A **500**, 485; Haidenbauer J and Meissner U-G 2005 Phys. Rev. C **72**, 044005.
- [12] Polinder H, Haidenbauer J and Meissner U-G 2006 Nucl. Phys. A **779**, 244; Haidenbauer J, Petschauer S, Kaiser N, Meissner U-G, Nogga A and Weise W 2013 Nucl. Phys. A **915** 24
- [13] *Nuclear Methods and Nuclear Equation of State* (International Review of Nuclear Physics, Vol 8. Edited by Marcello Baldo. World Scientific Publishing Company 1999).
- [14] Hulse R A and Taylor J H 1975 Astrophys. J. Lett. **195** L51.
- [15] Champion D J *et al.*, 2008 Science **320** 1309; Freire P C C 2009 arXiv:0907.32199; arXiv:1011.5809.
- [16] Demorest P *et al.*, 2010 Nature **467** 1081.
- [17] Antoniadis J *et al.*, 2013 Science **340** 6131.
- [18] Bednarek I, Haensel P, Zdunik L, Bejger M and Mańka R 2012 Astron. and Astrophys. **157** 543.
- [19] Weissenborn S, Chatterjee D and Schaffner-Bielich J 2012 Phys. Rev. C **85**, 065802; 2014 Phys. Rev. C **90** (E) 019904.
- [20] Oertel M, Providência C, Gulminelli F and Raduta Ad R 2015 J. Phys. G. **42** 075202.
- [21] Maslov K A, Kolomeitsev E E and Voskresensky D N 2015 Phys. Lett. B **748** 369.
- [22] Takatsuka T *et al.*, 2002 Eur. Phys. J. A **13** 213; 2008 Prog. Theor. Phys. Suppl. **174** 80.
- [23] Vidaña I, Logoteta D, Providência C, Polls A and Bombaci I 2011 Eur. Phys. Lett. **94**11002.
- [24] Yamamoto Y, Furumoto T, Yasutake B and Rijken Th A 2013 Phys. Rev. C **88** 022801; 2014, Phys. Rev. C **90** 045805.
- [25] Lonardoni D, Lovato A, Gandolfi S and Pederiva F 2015 Phys. Rev. Lett. **114** 092301.
- [26] Zdunik J L and Haensel P 2013 Astron. and Astrophys. **551** A61.
- [27] Özel F Psaltis D, Ransom S, Demorest P and Alford M 2010 Astrophys. J. Lett. **724** L199.
- [28] Weissenborn S, Sagert I, Pagliara G, Hempel M, Schaeffner-Bielich J 2011 Astrophys. J. Lett. **740** L14.
- [29] Klähn T, Blaschke D and Lastowiecki D 2013 Phys. Rev. D **88** 085001.
- [30] Bonanno L and Sedrakian A 2012 Astron. and Astrophys. **539** 416.
- [31] Lastowiecki R, Blaschke D, Grigorian H and Typel S 2012 Acta Phys. Polon. Suppl., **5** 535.
- [32] Hashimoto O and Tamura H 2006, Prog. Part. Nucl. Phys. **57** 564.
- [33] Keil W and Janka H-Th 1996 Astron. and Astrophys. **296** 145; Bombaci I 1996 Astron. and Astrophys. **305** 871; Prakash M, Bombaci I, Prakash M, Ellis P J, Knorren R and Lattimer J M 1997 Phys. Rep. **280** 1.
- [34] Vidaña I, Bombaci I, Polls I and Ramos A 2003 Astron. and Astrophys. **399** 687.
- [35] Burgio G F, Schulze H-J and Li A 2011 Phys. Rev. C **83** 025804.
- [36] Lattimer J M, Pethick C J , Prakash M and Haensel P 1991 Phys. Rev. Lett. **66** 2701.
- [37] Balberg S and Barnea N 1998 Phys. Rev. C **57** 409; Takatsuka T and Tamagaki R 1999 Prog. Theor. Phys. **102** 1043; Takatsuka T, Nishizaki S, Yamamoto Y and Tamagaki R 2000 Prog. Theor. Phys. **105** 179; Takatsuka T, Nishizaki S, Yamamoto Y and Tamagaki R 2002 Prog. Theor. Phys. Suppl. **146** 279; Vidaña I and Tolós 2004 Phys. Rev. C **70** 028802; Zhou X-R, Schulze H-J, Pan F and Drayer J P 2005 Phys. Rev. Lett. **95** 051101; Wag Y N and Shen H 2010 Phys. Rev. C **81** 025801.
- [38] Lindblom L 1986 Astrophys. J. **303** 146; Friedman J L, Ipser J R and Parker L 1986 Astrophys. J. **304** 115.
- [39] Lindblom L 1995 Astrophys. J. **438** 265.
- [40] Anderson N 1998 Astrophys. J. **502** 708; Friedman J L and Morsink S M 1998 Astrophys. J. **502** 714.
- [41] Chandrasekhar S 1970 Phys. Rev. Lett. **24** 611; Friedman J L and Schutz B F 1978 Astrophys. J. **221** 937; Friedman J L and Schutz B F 1978 Astrophys. J. **222** 281; Friedman J L 1978 Commun. Math. Phys. **62** 247.
- [42] Langer W D and Cameron A G W 1969 Astrophys. Space Sci. **5** 213; Jones P B Proc. R. Soc. A 1971 **323** 111; Levin Y 1999 Astrophys. J. **517** 328; Jones P B 2001 Phys. Rev. Lett. **86** 1384; Jones P B 2001 Phys. Rev. D **64** 084003; Lindblom L and Owen B J 2002 Phys. Rev. D **65** 0653006; Haensel P, Levenfish K P and Yakovlev D G 2002 Astron. and Astrophys. **381** 1080; van Dalen E N E and Dieperink A E 2004 Phys. Rev. C **69** 025802; Chatterjee D and Bandyopadhyay D 2006 Phys. Rev. D **74** 023003; Bondarescu R, Teukolsky S A and Wasserman I 2007 Phys. Rev. D **76** 064019; Chatterjee D and Bandyopadhyay D 2008 Astrophys. J. **680** 686; Gusakov M E and Kantor E M 2008 Phys. Rev. D **78** 083006; Sinha M and Bandyopadhyay D 2009 Phys. Rev. D **79** 123001; Patruno A 2010 Astrophys. J. **722** 909; Jha T K, Mishra H and Sreekanth V 2010 Phys. Rev. C **82** 025803.
- [43] Vidaña I and Albertus C 2015 in preparation.