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The neutron in astrophysics*

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Abstract. The discovery of the neutron 50 years ago has had an immense impact on our understanding of nuclear transmutations and in harnessing the vast amounts of energy released in nuclear fission and fusion reactions under terrestrial conditions. It is being increasingly realized that the neutron also holds the key to the solution of several astrophysical enigmas. To mention a few, the cosmic buildup of the heaviest elements such as gold or uranium involves rapid production and capture of neutrons; formation of rare isotopes (such as technetium seen in some stars) again involves neutrons; and the initial helium produced in the big bang results from an interesting interplay of the physical parameters characterizing the neutron and the ambient conditions in the early universe such as density and temperature. This primordial helium content in turn determines the subsequent evolution of stars and galaxies. Again in the context of recent theories of unification of fundamental forces, there seems to be an interesting connection between the neutron electric dipole moment and the universal photon-to-baryon ratio, which is another critical quantity determining the evolution of the universe. Finally the neutron has given us profound insight into how matter might behave at the end point of stellar evolution, which is of key importance in understanding the behaviour of pulsars for instance.

Key words : neutron—early universe—element formation—stellar evolution

Introduction

The neutron was discovered 50 years ago by Chadwick (1932, 1933) and was the culmination of efforts to understand the nature of the mysterious beryllium rays observed by Bethe and Becker when beryllium was bombarded by α-particles. The radiation was found to be too penetrating and energetic for it to consist of gamma rays and had the curious property of knocking off protons from hydrogen-rich materials like paraffin—something gamma rays cannot do. Chadwick first showed that all the properties of the rays are consistent with those of a neutral particle

*To commemorate the 50th anniversary of the discovery of the neutron.

slightly heavier than a proton, the reaction being $^4Be_9 + \alpha \rightarrow {}^6C_{12} + n$. It was soon shown that the neutron was a constituent along with protons of all nuclei except hydrogen. The simplest nuclear system known is the deuteron, the nucleus of deuterium being a triplet state of the two nucleon system comprising a proton and a neutron. The binding energy of the deuteron is quite small ($\sim$ 2.26 MeV) and that it would easily be photodisintegrated into a proton and neutron by a gamma ray photon of the appropriate energy was soon demonstrated by Chadwick & Goldhaber (1933). Again the cross-section for the capture of a neutron by a proton to form a deuteron and γ-ray is quite high (*i.e.* $n + p \rightarrow D + \gamma$). We shall see later that both these facts are very important for the production of deuterium and helium in the early stages of the big bang. As remarked the deuteron is rather loosely bound and it turns out that a change in the coupling strength of the nuclear forces by a few per cent may be sufficient to unbind it. In fact a decrease in the coupling constant of the strong interaction by hardly 5% is sufficient for the binding energy of the deuteron to vanish and for it to cease to exist in nature as a stable nucleus. Such a decrease would have had drastic implications in astrophysics as no deuterium would have been synthesized in the early big bang and consequently no helium; and moreover the p–p reaction in stars could not have taken place. (As it is, the first step in the proton-proton chain, *i.e.* $p + p \rightarrow D + e^+ + \nu$, is such a slow reaction, the cross-section is only $\sim 10^{-47}$ cm^2, *i.e.* 10 sheds, 1 shed = 10^{-48} cm^2 as opposed to a barn which is 10^{-24} cm^2, that it is not yet been observed under laboratory conditions, the halflife being a few billion years.) Consequently the whole universe would have a composition of 100% hydrogen and nothing else. Again nuclear forces are spin (isospin) dependent, the potential energy of an n-nucleon system may be written as

$$V(n) = \tfrac{1}{2} \sum_{i \neq j} V_{ij} = \tfrac{1}{2} V \sum_{i \neq j} \bar{t}_i . \bar{t}_j,$$

where t_i is the isospin of the ith particle and V includes dependence on other parameters.

Thus $$V(n) = \tfrac{1}{2} V [I(I + 1) - nt(t + 1)],$$

I being the total isospin of the system and t the nucleon isospin. This potential is seen to be attractive (V negative) for the deuteron which has ($I = 0$), and repulsive for the dineutron (*i.e.* a nucleus composed of two neutrons) and diproton (two-proton nucleus) which have $I = 1$. This repulsive part of the potential is the reason why one does not have a nucleus in nature consisting of just two protons or two neutrons. However it turns out that if the nuclear attraction coupling constant were only about 2% stronger, then this would have been enough to bind the diproton (*i.e.* ^{2}He). If the diproton were bound, then the basic nuclear interaction in main sequence stars would no longer have been the extremely slow $p + p \rightarrow D + e^+ + \nu$ reaction but the rapid strong interaction $p + p \rightarrow {}^2He + \gamma$. The subsequent spontaneous decay of ^{2}He, *i.e.* $^2He \rightarrow D + e^+ + \nu$ would not have decreased the rate of burning of the hydrogen so that the process would have manifested itself as an instantaneous explosive release of energy making all stars unstable. This catastrophic conversion would also have occurred in the early universe at such a rapid rate that practically no hydrogen would have been left. That this has not happened is evidence for the remarkable constancy of the strong interaction constant.

On a cosmic scale, it turns out that neutron induced reactions are responsible for the astrophysical production of a wide variety of elements and a large number of their various isotopes. Gamow (1948) pointed out that successive neutron captures would produce an inverse relationship between abundance and cross-section. This correlation covers almost the entire periodic table and particularly the closed shell nuclei immediately suggesting models of primeval nucleosynthesis in the early universe. Again it was Gamow (1946) who proposed that the early stages of the universal expansion could have been hot and dense enough for the production of the heavy elements. He considered an initial all-neutron state, in which the universe began as a dense hot neutron ball. Now it is known that this picture was grossly oversimplified and we shall see in the next section that it is now believed that only helium and a trace amount of lithium could have been produced in the early universe. Elements heavier than the iron group are however produced in the later stages of stellar evolution essentially by the processes of rapid and slow neutron capture (the socalled r- and s-processes respectively). We shall review these processes in later sections. The r- and s-processes finally solved the riddle of the heavy element catastrophe which beset the earlier equilibrium theory (e-process) of the origin of elements which was proposed soon after the discovery of the neutron (Sterne 1973). Finally we shall deal with other important roles for the neutron in astrophysics, such as in giving us an insight into the nature of the endpoint of stellar evolution; and more recent applications involving neutron oscillations and the observed excess low energy antiproton background (Sivaram & Krishan 1982, 1983a); and interesting relation between a possible neutron electric dipole moment and the photon-to-baryon ratio in the universe (Ellis *et al.* 1981).

In the next section we describe in some detail how the initial helium abundance in the early stages of the big bang depends crucially on the physical parameters characterizing the neutron.

2. The neutron and the big bang

There is by now a well known argument about the cosmic abundance of helium being too large to be easily explicable in terms of nucleosynthesis in stars (*e.g.* Hoyle & Tayler 1964; Cameron 1965). The luminosity-to-mass ratio L/M of our galaxy is about a tenth the solar ratio or 0.2 erg gm^{-1} s^{-1}. If the luminosity of the galaxy has remained constant during the last 10^{10} years, then the energy produced per nucleon is about 0.06 MeV. On the contrary, the thermonuclear fusion of hydrogen into helium releases about 6 MeV per nucleon, thus implying that not more than about 1% of the nucleons in our galaxy could have been fused into helium (or heavier nuclei) by the usual stellar processes. Estimates of the helium abundance at present vary but apparently the cosmic abundance of helium by mass is considerably greater than 1% and is generally believed to lie between 20 and 25%. It is one of the remarkable features of the standard big bang model that it naturally accounts for this figure for the helium abundance with very few input parameters such as the neutron-proton mass difference. It is of course quite possible that the requisite amount of helium could have been synthesized in earlier more luminous epochs of the galaxies (after all, we assumed that the galaxy had the same luminosity throughout its history of 10^{10} years) or in some supermassive objects present in an earlier epoch just after the recombination era. Another curious coincidence is that the total energy released in the synthesis of about 25% of the nucleons into helium (which is about 10^{73} erg) would, if thermalized, account for the present 2.7 K microwave background. This has formed the basis of certain alternative ways of generating both the microwave background and helium abundance through the evolution of primordial supermassive objects (Rees 1978; Sivaram 1982, 1983d). However the big bang accounts vary naturally for both these large scale features of the universe with very little input parameters as noted above, unlike other scenarios which involve some degree of arbitrariness and have some inherent difficulties (Sivaram 1983d) such as circumventing the very slow p–p reaction, other reactions already involve heavier isotopes which would be hard to account for without the requisite helium. Gamow argued that although the early hot dense period of cosmic expansion was much briefer than the lifetime of a star, there was a large number of free neutrons present at that time so that the heavy elements could be rapidly built up by successive neutron captures, starting with $n + p \to d + \gamma$. The abundances of the elements would then be correlated with their neutron capture cross sections in rough agreement with observation and the necessity of avoiding overproduction of helium required the presence of blackbody radiation with an estimated present temperature of 5 K. (Alpher & Herman 1950). However a major snag here was that there are no stable nuclei with atomic weight $A = 5$ or $A = 1$—both He^5 and Be^8 are extremely unstable with half lifes $\sim 10^{-21}$ s—so that it is difficult to build up elements heavier than helium by $p - \alpha$, $n - \alpha$, or $\alpha - \alpha$ reactions. But in stars which have converted all the hydrogen to helium in their dense cores it is possible to bridge these gaps at $A = 5$ or 8 by the production of Be^8 in $\alpha - \alpha$ collisions followed by production of C^{12} in $\alpha - Be^8$ with a reaction time scale (owing to high densities) shorter than the decay time of Be^8. The formation of C^{12} is considerably helped by the presence of a resonant excited level (as predicted by Hoyle) of C^{12}, at 7.64 MeV close to the energy of the $\alpha - Be^8$ system. Moreover the expansion rate of the universe and the consequent diminishing of temperature and dilution of matter prevent all the nucleons from being converted to helium. Hayashi (1950) pointed out that unlike in the picture of Gamow it would not be posssible to have a high temperature and nothing but radiation and neutrons. At a temperature T particle-antiparticle pairs would be created from radiation by reactions of the type $\gamma + \gamma \to p + \bar{p}$, so long as the mass of a typical particle p satisfies $kT \geqslant m_p c^2$. In other words both neutrons and protons would be present and the reactions converting neutrons into protons would produce a neutron-to-proton (n/p) number ratio which at a high temperature would have the form: $(n/p) = \exp(-\Delta mc^2/kT)$, where

$$\Delta m = \text{neutron-proton mass difference} = m_n - m_p$$
$$\simeq 2.5\, m_e = 1.293 \text{ MeV}.$$

This relative abudance is determined by detailed balance and a solution of the Boltzmann equation in a Friedmann model universe, shows that particle number for any particle of mass m, is proportional to $\exp(-mc^2/kT)$. At a sufficiently high temperature neutrons and protons are maintained in numerical equilibrium by the the weak interactions:

$$n + \nu \rightleftharpoons p + e^-; \quad n + e^+ \rightleftharpoons p + \bar{\nu}; \quad n \rightleftharpoons p + e^- + \bar{\nu}.$$

The rate of these weak interactions can be estimated as $R_W = n\sigma V$, where n the number density for all different species at high temperatures $\sim (kT/\hbar c)^3$. The cross-section for processes involving the neutrinos and neutrons and protons and electrons is given by weak interaction theory as $\sigma = G_F^2 (kT/\hbar^2 c^2)^2$, where G_F is the universal Fermi weak constant and one notes the dependence of the neutrino cross-section on the energy, *i.e.* on temperature squared. Then with $V \sim c$, R is: $R_W \simeq G_F^2 \hbar^{-7} \times (KT)^5 c^{-6}$ s^{-1}. As is well known the expansion rate of the universe is governed by $H = \dot{R}/R = (8\pi G\rho/3)^{1/2} = (8\pi G N_1/3)^{1/2}\ T^2$, where N_1 is the number of particle species present. At a sufficiently high temperature $T > 10^{11}$ K, it is seen that $R_W \gg H$, so that the neutrons and protons are kept in equilibrium. However as can be seen from above, the weak interaction rate goes as T^5 whereas the expansion rate goes as T^2, so that as the temperature drops with expansion of the universe, the rate of weak interactions maintaining equilibrium falls off much faster than the expansion, and below a particular temperature, the so called *freeze-out* temperature T_F, the protons and neutrons are no longer kept in equilibrium by the above weak interactions and the value of n/p is effectively frozen at the temperature T_F. T_F is determined by the value of T when the weak interaction rate and the rate of expansion become comparable

$$kT_F \sim G^{1/6} G_F^{-2/3} (\hbar^{11} c^{-5})^{1/6}.$$

Below this temperature the expansion rate dominates and the value n/p changes slowly owing to the decay of the neutrons. More rigourously, one has the differential equation for the ratio X_n of neutrons to all nucleons

$$-\frac{dX_n}{dt} = \lambda(n \to p)\, X_n - \lambda(p \to n)\,(1 - X_n),$$

and for $kT \gg \Delta mc^2$, we have (Peebles 1966)

$$\lambda(n \to p) \simeq \lambda(p \to n) \simeq \text{const} \int_{-\infty}^{\infty} q^4\, dq\, (1 + e^{-q/kT})^{-1}/(1 + e^{q/kT})$$

$$= 0.36\, (T/10^{10})^5 \text{ s}^{-1}, \ (q \text{ being the momentum}).$$

Rate of production of neutrons should be compared to the expansion rate for the universe

$$t = 1.09\, (T/10^{10})^{-2} \text{ s}.$$

Product $\lambda t > 10$, for $T \geqslant 3 \times 10^{10}$ K, so for these temperatures the equilibrium solution is

$$X_n \simeq \frac{\lambda(p \to n)}{\lambda(p \to n) + \lambda(n \to p)}$$

Thus the neutron abundance for $T \geqslant 3 \times 10^{10}$ K is $X_n \simeq [1 + \exp\,(\Delta mc^2/kT)]^{-1}$. The neutron abundance starts at $X_n \simeq \frac{1}{2}$ at very early times and drops slowly as the temperature falls, *i.e.* when $T \gg \Delta mc^2/k$, we have $n/p \sim 1$ but as T falls to near about Δm, it becomes more likely that the less massive proton is produced and n/p slips below unity. Once T drops to the freeze out temperature $T_F \sim 2 \times 10^9$K, the rates of the reactions $n + \nu_e \leftrightarrows p + e^-$, $n + e^+ \leftrightarrows p + \bar{\nu}_e$ and $p + e + \bar{\nu}_e \rightarrow n$, become negligible and the only remaining reaction is the free neutron decay $n \rightarrow p + e^- + \bar{\nu}_e$. The particular values of the dimensionless weak and gravitational interaction strengths, G and G_F, lead to a value of T_F which results in a finite $(n/p)_F$. The temperature falls to T_F after $t_F \sim 1$ s and thereafter some slight free neutron decay ($\lambda^{-1}(n \rightarrow p + e^- + \nu_e) = 1013$ s) will occur before nuclear reactions proceed rapidly at $T_N \leqslant 10^9$ K. Thus the neutron abundance from the time when $T_F \simeq 2 \times 10^9$ K to the start of nucleosynthesis is given by

$$X_n(t) = X_n^{(0)} \exp\left[-\frac{t\,(\mathrm{sec})}{1013}\right], \text{ where } X_n^{(0)} = (n/p)_F \simeq 0.16.$$

The first nuclear reaction to occur between the protons and neutrons is the radiative capture reaction $p + n \rightarrow D + \gamma$ which is unable to build up a deuterium abundance until T falls to about 10^9K. Because till then, the photonuclear dissociation reaction $D + \gamma \rightarrow p + n$ is equally rapid, *i.e.* the deuterons are dissociated as fast as they are produced. The cross-section for the capturere action is

$$\sigma_{n\gamma} = 2\pi\alpha(\mu_p - \mu_n)^2 \left(\frac{B_d}{Mc^2}\right)^{5/2}\left[\left(\frac{\hbar}{\mu_d B_d}\right)^{1/2} - S\right]^2,$$

where μ_p and μ_n are the proton and neutron magnetic moments, B_d is the binding energy of deuterium, S is the singlet $n - p$ scattering length, $(\hbar/\mu_d\, B_d)^{1/2}$ is the effective nuclear size. With this cross-section, the rate of deuterium production per free neutron is

$$\begin{aligned}\lambda_d &= [4.5 \times 10^{-20}\ \mathrm{cm^3\ s^{-1}}]\, n_p \\ &= 27\left(\frac{R}{10^{-9}R_0}\right)^{-3}\left(\frac{\rho_{N0}}{10^{-30}\ \mathrm{g/cc}}\right)X_p\ \mathrm{s^{-1}}\end{aligned}$$

and this rate is so much faster than the corresponding expansion rate t^{-1} that deuterons will appear with the equilibrium abundance

$$X_d = \frac{3}{\sqrt{2}} X_p\, X_n \hbar^3 m_N\, (2\pi m_N kT)^{-3/2} \exp\,(B_d/kT).$$

No appreciable quantity of H^3, He^3, He^4 or heavier nuclei can be formed until this equilibrium deuterium abundance is high enough to allow the sequence of two-body reactions such as :

$$D + D \rightarrow He^3 + n \rightarrow He^3 + p;\quad H^3 + D \leftrightarrows He^4 + n,$$

or

$$p + D \rightarrow He^{3-} + \gamma,\quad n + D \rightarrow H^3 + \gamma,$$

$$D + D \rightarrow He^4 + \gamma,\quad n + He^3 \rightarrow He^4 + \gamma, \text{ etc.}$$

The full network of these reactions can be integrated numerically (Peebles 1966; Wagner *et al.* 1967). The low binding energy of deuterium (and consequently its easy photodissociation) serves as a bottleneck which delays formation of more complex nuclei until T drops to just below 10^9 K. Once nucleosynthesis begins, it proceeds very rapidly and the bulk of the material ends up as a mixture of hydrogen and helium-4. It is not in fact possible to produce appreciable quantities of elements heavier than He^4, because as mentioned above the lack of stable nuclides at $A = 5$ or $A = 8$ impedes further nucleosynthesis via $n - \alpha$, $p - \alpha$ or $\alpha - \alpha$ collisions while the Coulomb barrier in the reactions $He^4 + H^3 \rightarrow Li^7 + \gamma$, and $He^4 + He^3 \rightarrow Be^7 + \gamma$ prevents them from having competing rates with $p + H^3 \rightarrow He^4 + \gamma$ or $n + He^3 \rightarrow He^4 + \gamma$. Thus the effect of the primordial nuclear reactions is to very rapidly incorporate all available neutrons into He^4 nuclei which have the largest binding energies of all nuclei with $A < 5$. Only traces of heavier elements, chiefly lithium-7 are synthesized. The effect of nucleosynthesis on the nuteron-proton ratio is that by "turning off" the decay of free neutrons it freezes this ratio at the value it had just before the onset of nucleosynthesis, that is at the quantity $X_n(0)$. After the nucleosynthesis is over we have effectively nothing but free protons and nuclei of He^4 so the fraction of neutrons to all nucleons is one half the fraction of all nucleons that are bound in helium, or one half the abundance by weight of helium. Thus the abundance of the cosmologically produced helium is simply: $Y \equiv 2X_n$, just before nucleosynthesis or 25%. It is worth noting that the helium abundance is determined principally by the neutron-proton ratio at T_F which depends very sensitively on T_E. The crucial role is played by the neutron-proton mass difference $\Delta mc^2 = 1.3$ MeV. Any change in this or a different Δm would give a rather different helium abundance, either no helium at all or too much helium. The neutron-proton mass difference still remains unexplained in the context of particle physics theories, some of the theories even giving the wrong sign for the difference ! But here the fact that a reasonable amount of helium is formed for the later evolution of main sequence stars is crucially dependent on the mass difference having this precise value and being of the right sign. The neutron lifetime also plays a role in determining the final helium abundance at the end of the nucleosynthesis and it is clear that the properties of neutron decay are crucial to a proper understanding of primeval element production. When Gamow proposed the theory initially, the neutron halflife was believed to be 14 min. Subsequent measurements have given the values 10.6 min (Christensen *et al.* 1972) and 10.8 min (Byrne *et al.* 1980). Wilkinson (1981) gets 10.6 min. A change in the neutron halflife by about half a minute—which is not ruled out—would result in a change in the helium content by about 4% of its value with a shorter halflife corresponding to a lower helium content. The neutron halflife is thus critically important as far as cosmology is concerned. It is also important to realize that the helium abundance is relatively insensitive to the matter density ($\propto \log \rho_M$), chiefly because at the epoch of synthesis the density is totally dominated by the radiation content, $\rho_\gamma > 10^6 \rho_M$ and so the matter plays only a second order role in determining final abundance of ^4He. Also none of the helium is destroyed once made thereby avoiding a dependance on ρ_M once again. In contrast, deuterium abundance is extremely sensitive to ρ_M with $X_D \propto \rho_M^{-1.3}$. The cross-sections for reactions converting D and He^3 into He^4 increase with ρ_M so that, the D and He^3 abundances drop while He^4 abundance increases. To obtain agreement of the calculated abundances of D, He^4 and Li^7 with observations, one must have a baryon-to-photon ratio of $n_B/n_\gamma \sim 5 \times 10^{-10}$. Before the notion of the possibility of neutrino rest mass became popular, the rather large value of observed D abundance in the universe, X (D) $\sim 10^{-5}$ was taken as evidence for a low ρ_M and consequently an open universe (Schramm 1977). A higher density would imply a lower D abundance. More rapid expansion rates (like in models of varying G, or due to presence of more species of noninteracting relativistic particles or massive particles with very weak coupling) would imply earlier freezing out of n/p ratio with a larger value and therefore a greater helium abundance We have thus seen that the neutron plays a key role in substantiating the claims of the big bang model of the universe. It is interesting that the neutron was also invoked in the steady state theory of Gold & Hoyle (1959). In this theory, the large scale properties of the universe do not change with time and continuous creation of matter is required to keep the mean density of the universe constant as it expands. In the original version, the created matter was assumed to be hydrogen. However Gold & Hoyle suggested that the matter might be created as neutrons rather than as hydrogen. This would not affect nucleosynthesis but the decay of the neutrons would produce a very hot intergalactic medium at a temperature of 10^9 K in pressure balance with matter inside galaxies. This hot medium would become thermally unstable and form selfgravitating condensations and then new galaxies needed by the steady state theory. Very energetic particles would be produced when the instability developed. Thus the continuous creation of neutrons would play a crucial role in the development of structure in the steady state universe. Unfortunately for the theory, it soon became clear that the bremsstrahlung radiation from the hot gas was much more than what was received at the earth (Gould & Burbidge 1963; Field 1972). A recent version of the 'cold' big bang with neutrons playing a crucial role is due to Hogan (1982).

3. Neutrons and the formation of heavy elements in stars

We have seen the crucial role played by the neutron in the primordial nucleosynthesis of helium. We also saw that apart from a trace production of Li^7 no other heavier elements are formed in the big bang. All the other heavier nuclei must therefore have been made in nuclear processes occurring during the course of stellar evolution. In fact the large scale features of the abundance curve of the nuclei found in nature when plotted against atomic weight suggest that the most likely source of most of the nuclei is a sequence of more or less discrete processes, *i.e.* groups of nuclear processes or reactions operating within stars. As first listed in detail by Burbidge *et al.* (1957 = B^2FH) these processes are now called hydrogen, helium, carbon, neon, oxygen and silicon burning and for the still heavier elements the s and r processes. As this review is primarily concerned with outlining the role played by the neutron in astrophysics, we shall be chiefly concerned in describing the s and r processes. However for completeness we shall briefly describe the other processes in the order in which they occur in stellar evolution. Hydrogen burning is the fusion of four protons to form a He^4 nucleus either via the slow p–p chain of nuclear reactions or in catalytic cycles of nuclear reactions involving carbon, nitrogen and oxygen. Most of the power required to generate the tremendous luminosity of stars for the greater part of their lives comes from this fusion of hydrogen into helium either in the cores of the stars or at later stages in thin, spherical shells surrounding the helium cores, which are the ashes of the hydrogen burning. When the helium core of an evolved star becomes sufficiently hot and dense while contracting under its gravity, helium burning can be ignited in the core, the helium reacting with itself to produce carbon and oxygen. Depending on the star's mass, the star may then either be completely disintegrated by igniting carbon under highly degenerate explosive conditions or, if rather massive, burn carbon under hydrostatic conditions. If it survives carbon burning, the star can pass through a brief stage in which neon nuclei produced earlier in carbon burning can photodisintegrate to $O_{16} + \alpha$. The α-particles released can then add on to undissociated neon to build Mg_{24}. This is called neon burning and leaves a core of chiefly O_{16} and Mg_{24}. This entails further contraction of the stellar core, raising its temperature until the oxygen nuclei fuse with each other releasing α's, protons and neutrons building a range of elements extending from silicon to scandium. These successive stellar evolutionary stages occur on decreasing timescales and oxygen burning may merge in to the next stage of evolution, *i.e.* silicon burning which occurs at temperatures of several billion degrees, on a very short timescale. Protons, neutrons and α-particles are ejected by photodisintegeration from the most abundant nuclei present and made available for building nuclei up to and slightly beyond the iron peak in the abundance curve. This happens at temperatures in excess of 3×10^9 K. Silicon burning will result in significant release of energy if the temperature is $< 5 \times 10^9$ K. For higher temperatures the very intense photon flux maintains large number of α-particles and free nucleons. This situation leads to a significant reduction in the total binding energy and final product of the reactions which are not endoenergic are $Fe^{54} + 2p$. If Si burning takes place slowly (at $T < 3 \times 10^9$ K) β-decays will be important and Fe^{56} can be directly produced. At somewhat higher temperatures, silicon burning is simply $2\,Si^{28} \rightarrow Ni^{56}$ which is exoergic by 11 MeV, with Ni^{56} decaying to Fe^{56}. At iron, the binding energy per nucleon is a maximum, and consequently and beyond the iron peak it consumes more energy to photodissociate a nucleus than can be gained by adding nucleons to the nucleus. This is the reason why the processes invoked by earlier workers, such as the equilibrium or e-process (Hoyle 1946) could not account for elements beyond the iron group as such processes favour the production of elements with the largest binding energy per nucleon as is apparent from a study of the appropriate Boltzmann and Saha equations. It thus becomes clear that elements heavier than iron must mainly be the products of nucleosynthesis by neutron capture. B^2FH also pointed out that at 8×10^9 K, statistical equilibrium favours breakup of Fe^{56} as $Fe^{56} \rightarrow 13\alpha + 4$ neutrons. The conversion of 1 gm of iron to helium requires an energy supply of 1.6×10^{18} erg whereas the total thermal energy at that temperature is only 3×10^{17} erg g^{-1}. Thus the energy for the conversion must come from the gravitational energy released by the shrinking of the star and this must be very large, implying an implosion of the central regions in a time scale $\sim 1/5$ s and according to B^2FH this catastrophic implosion triggers the outburst of a supernova. It is to be noted that each disintegeration is accompanied by four neutrons. A lot of work has been done on nucleosynthesis at both the presupernova and supernova stages (Weaver & Wooley 1980). Confirmation of the essential correctness of the above scenarios has come from recent studies of optical filaments in the 300 yr-old supernova remnant, Cas A, which are strongly overabundant in sulphur, argon and calcium relative to lighter elements, just the expected products of oxygen burning (Kirshner & Blair 1980). Also the x-ray spectrum of the remnant of Tycho's supernova exhibits similar strong enhancements of silicon, sulphur, argon and calcium (Becker *et al.* 1980). These objects apparently exploded prior to iron formation. A neutron-rich environment is unavoidable for the formation of elements and isotopes heavier than iron. For one thing we have higher abundances of neutron-rich isotopes than of lighter isotopes. Secondly as already noted there is the impossibility of building up the heaviest elements by charged particle nuclear reactions because the temperatures required would need to be so high that photodissociation of nuclei would be more important than thermonuclear fusion. Two major processes have been invoked by which neutron capture reactions can build up heavier elements. These are the r and s processes; r stands for rapid capture of neutrons and s for slow capture of neutrons, the processes being differentiated as their names suggest by the different timescales of the reactions. Both the r and the s processes involve neutron additions to nuclei starting with seed nuclei, which are primarily Fe^{56} with minor contributions from other iron peak species. The neutron flux required for the r-process is so high that the nuclei subjected to the irradiation are so enriched with neutrons that their halflifes are of order 1 s whereas the s-process species with halflives of about 10–100 yr apparently had time to decay. The r-process timescale is based primarily on the mass of the peak abundances resulting from the closed neutron shells, these peaks occuring at atomic masses which are lower than the masses of stable nuclei with closed neutron shells by about ten atomic mass units. The s-process critical timescale indicators are the nuclei associated with Se 79 and Sm 151. The abundances of Kr 82 and Gd 154 indicate that these decays did take place and as the laboratory halflives of the unstable nuclei are 10^5 yr and 10^2 yr respectively, it is apparent that the neutrons were added over long timescale. Apparently the s-process occurs in a stellar core whereas an explosive environment is a more appropriate site for r-process. Clayton *et al.* (1961) showed that a single neutron exposure produces a distribution of abundance unlike that observed in carbonaceous chondrites. B^2FH (1957) pointed out that the product of the neutron capture cross-section σ_c and the s-process abundance N_s is nearly constant between closed shells for which σ_c is small. Clayton *et al.* found that $\sigma_c N_s$ had a peak at the closed shell, but near the closed shell it either decreased sharply or increased sharply depending on whether the seed nuclei had captured enough neutrons to produce nuclei lighter or heavier than the closed shell nucleus. The solar system abundances have $\sigma_c N_s$ decrease at each closed shell but the heaviest nuclei (*e.g.* 208 Pb) certainly have had significant contributions from the s-process. Seeger *et al.* (1965) showed that an exponential distribution of exposures produces reasonable fits to observed abundances. It now seems likely that the bulk of the s-process distribution is produced in individual stars and the most probable birthplaces are the inner regions of red giants. Many red giants show unusual surface compositions of the s-product elements. Evidence that composition anomalies do not merely reflect differences in a star's initial material content is illustrated by the remarkable behaviour of the star FG Sagittae.

Since 1967 the surface abundance of the s-process elements in this star has increased to a value approximately 25 times the solar abundance, although prior to 1965 these elements were not substantially enhanced (Langer *et al.* 1974). Moreover the discovery of radioactive technetium (with the longest halflife of only 2×10^5 yr) in some red giant stars is a strong indication that s-process synthesis does occur during a star's evolution in the red giant phase*. Through mass loss, the red giant stars therefore can contribute to the heavy element enrichment of the interstellar medium. An early suggestion for the principal neutron source within the star to operate the s-process was the reaction $^{13}C + \alpha \rightarrow {}^{16}O + n$. But the fact that ^{14}N is the dominant product of C — N — O hydrogen burning would pose problems, as ^{14}N can capture neutrons with a large cross-section through the reaction $^{14}N + n \rightarrow {}^{14}C + p$. Incidentally this particular well know reaction is the key to radioactive carbon-dating as neutrons from cosmic rays interact with nitrogen in the earth's atmosphere, to produce trace amounts of C^{14} which forming part of CO_2 enter into organic metabolism. C^{14} decays in dead organic matter which is no longer in metabolic equilibrium with the atmosphere with a halflife of 5700 yr. The proton produced in the above reaction would again generate another N^{14}. However the status of $^{13}C(\alpha, n)\ O^{16}$ as an important neutron source was reestablished following Schwarzschild & Harm (1967) who noted that there would be an instability during double shell burning, *i.e.* during the second red-giant phase when there is an inert carbon-oxygen core surrounded by a helium burning shell, an intershell region largely of helium and further out a hydrogen burning shell. This could take fresh hydrogen from the outer boundary of the hydrogen burning shell down to the region where ^{12}C is being formed by the triple α-process in the helium burning shell. Thus by admitting about one proton to every C^{12}, C^{13} would be made without any significant N^{14} being formed. The C^{13} thus produced will be carried inward by convection until it is destroyed by the reaction $C^{13}(\alpha, n)O^{16}$ at the prevailing higher temperatures near the helium-burning shell. Net result is the liberation of one neutron for each proton mixed into the intershell region. The s-process elements can then be built up by successive absorption of these neutrons by the iron peak nuclei (Cameron & Fowler 1971). The convective shell through which the protons are added is formed during thermal relaxation oscillations of the helium shell source. These thermal pulses occur because the helium shell source is so thin that the addition of thermal energy causes the temperatuie to rise instead of causing the star to expand and cool as happens on the main sequence. If mixing occurs for one flash it will probably occur for a few subsequent flashes also as the structure of the star changes only slightly from one cycle to the next. Because the relaxation oscillations produce a distribution of $\sigma_C N_S$ similar to the solar system distribution, within a single star these oscillations have received the bulk of the theoretical analysis in the last decade (Iben 1975a, b; Iben & Truran 1978). The question is whether the envelope convection at its deepest penetration will reach through the hydrogen burning shell and into the intershell region, for in such a case helium, carbon as

*The element technetium, no. 43 in the periodic table, does not exist on earth as the halflife of all its isotopes is considerably shorter than the age of the earth; it was artificially produced on earth by bombarding nuclei with neutrons, much the same way as in the s-process.

well as any s-process elements previously formed would be transported outward to the surface accounting for chemical anomalies. At least for the first ten cycles of a 7 $M_\odot$ star it turns out that the intershell convective zone does not reach the burning shell. However for such a star there is another attractive alternative neutron source:

$$Ne^{22} + \alpha \rightarrow Mg^{25} + n.$$

The Ne^{22} is formed from the N^{14} in the intershell region by two successive alpha capture reactions during the flashes. The above neutron producing reaction proceeds at a significant rate during later cycles when the maximum flash temperatures are higher and this reaction would then yield the necessary neutrons for s-process synthesis. Some of the R stars, the CH stars and the Ba stars show abundance anomalies that are most likely a result of the He flash and subsequent processes. Apart from the favourite ^{22}Ne neutron source, there are several other helium-induced reactions that may in some situations be significant as neutron sources for subsequent s-process synthesis. Some of these are :

$$^{17}O(\alpha, n)\,^{20}Ne, \quad ^{18}O(\alpha, n)\,^{21}Ne, \quad ^{21}Ne(\alpha, n)\,^{24}Mg, \quad ^{26}Mg(\alpha, n)^{29}Si.$$

A summary of the equations underlying the s-process during the relaxation oscillations is now given. Let N_A be the abundance of a species with atomic mass A, $N_A(S)$ being the abundance remaining in the convective shell of mass M_C from the previous fiash. At maximum convective phase the total number of nuclei remaining from the previous cycle is $\gamma M_C N_A$. The abundance of A is changed by exposure to neutrons released by either the ^{22}Ne or ^{13}C source, that is activated during the peak of the cycle. The exposure to neutrons can be described by $\Delta\tau = \int N(t)\, V_{th}\, dt$, this being set by the number of neutrons released by the neutron source and the average cross-section of the matter in the shell. Nuclei of mass $A - 1$ capture a neutron to form species A which in turn is destroyed by capturing a neutron. The equation describing the process is :

$$\frac{dN_A}{dt} = \sigma_{A-1} N_{A-1} - \sigma_A N_A,$$

where the initial condition is :

$$N_A(0) = \gamma N_A(\Delta\tau) + (1 - \gamma)\, {}^{Ext}_{A}\, .$$

N_A^{xEt} is the number of nuclei added; $N_A(0)$, $N_A(\Delta\tau)$ are the abundances at the beginning and end of the pulse respectively. The process can be regarded as nearly continuous for small $\Delta\tau$ and $\gamma \approx 1$. The number $M_C\,(1 - \gamma)\, N_A$ of nuclei lost to the region are effectively destroyed. Thus

$$\frac{dN_A}{dt} = \sigma_{A-1} N_{A-1} - \sigma_A N_A - \lambda\,(N_A - N_A^{Ext}),\ \lambda = \frac{1 - \gamma}{\Delta\tau}.$$

Then approximately $(\sigma_A + \lambda)\, N_A = \sigma_{A-1} N_{A-1} + \lambda N_A^{Ext.}$

or $$\sigma_A N_A = \sigma_{A-1} N_{A-1}\,(1 + \lambda/\sigma_A)^{-1} + \lambda N_A^{Ext}\,(1 + \lambda/\sigma_A)^{-1}.$$

Repeated application for successive species gives

$$\sigma_A N_A = \sum_{A'=1}^{A} \lambda N_{A'}^{Ext} \prod_{A''=A}^{A} (1 + \lambda/\sigma_{A''})^{-1}.$$

For species beyond iron, it is necessary to include for N_A^{Ext}, only the iron peak elements. The resulting distribution is nearly exponential as was required by Clayton & Ward (1974). The distribution of $\sigma_A N_A$ with A is in good agreement with solar system abundances for $\lambda = 4 - 5$ mb. Again factors affecting halflives (such as degeneracy) must be taken into account. Quite generally if we assume a target of species A bombarded during a time t with a neutron flux ϕ, having a range R (in gm cm^{-2}) and a cross-section σ_A and species $(A + 1)$ decays with a halflife of $\tau_{1/2}$ or $\lambda = \ln^2/\tau_{1/2}$ then the number of active nuclei, $(A + 1, Z + 1)$, present is given by

$$\frac{dN_{A+1}}{dt} = \phi\sigma_A N_A - \lambda N_{A+1}, \text{ and } \frac{dN_A}{dt} = -\frac{dN_{A+1}}{dt}.$$

The solution to these equations (satisfying $N_A(0) = RA/M_A$ (A = shell area, M_A = mass of nucleus A), is given by

$$N_{A+1}(t) = \frac{RA\phi\sigma}{M_A(\lambda - \phi\sigma)} [\exp(-\phi\sigma t) - \exp(-\lambda t)].$$

In the s-process, neutron capture occurs on a 10^2–10^6 yr timescale and consequently all energetically possible beta decays are completed before additional neutron capture. Thus this process produces nuclides that are close to the valley of beta stability. The s-process is terminated at bismuth and polonium, because the subsequent isotopes produced are α-emitters of short halflife. This implies that s-process nucleosynthesis cannot be responsible for the formation of uranium, thorium or a number of other stable neutron rich isotopes. On the contrary several element (*e.g.* zirconium, barium and technetium) predicted to be produced by means of the s-process are observed to be overabundant in some red giants, confirming that the ideas outlined above are essentially correct. In a non-steady s-process, the neutron exposure during a single relaxation cycle, $\Delta\tau$, is a crucial parameter, for with species for which $\sigma_A\Delta\tau \gg 1$ the abundances during the cycle come quickly to a steady state and consequently it turns out that any species with $\sigma \gg 500$ mb will achieve a rough steady state during the oscillation. Among the species satisfying this condition are

$$^{134}\text{Cs}, {}^{151}\text{Sm}, {}^{154}\text{Eu}, {}^{163}\text{Dy}, {}^{170}\text{Tm}, {}^{182}\text{Ta and } {}^{185}\text{W}.$$

Once the neutron irradiation ceases, the decays all take place except for the longest lived species. The abundances of the decay product is obtained by adding the abundances of the unstable species and these of the end product, which both have steady state values during the active part of the relaxation oscillation. For the duration of the inactive period the abundance is

$$\sigma_{A-1,\, z+1} N_{A-1,\, z+1} = \frac{\sigma_{A-1,\, z+1} V_{th} N_n \tau_{1/2} + 1}{\sigma_{A-1, z} V_{th} N_n \tau_{1/2} + 1} \sigma_{A-2} N v_{-2}.$$

As a result of this 'decay' effect, some species normally considered to be falling on the path of the s-process can be bypassed. Some significant examples are ^{154}Gd, ^{134}Ba, ^{170}Yb and ^{160}Dy. The critical factors that lead to difficulties in the models for producing the heavier elements are the required strength of the neutron source and the neutron densities, *i.e.* it has to be explained how the bulk of the enrichment seems to have been produced at the rather low available neutron densities. We have already mentioned that s-process nucleosynthesis could not have formed elements like uranium or thorium. For the formation of these stable neutron-rich isotopes it has been suggested that r-process is responsible, and it is generally believed that all nuclei heavier than 209 Bi were synthesized by this process. The r-process requires the neutron capture rates to be much faster than those for beta-decay unlike the s-process, the rapid neutron capture must take place in a time scale t of about 10^{-5} s per capture. The nuclei capture neutrons via (n, γ) radiative reactions whose rates are fast compared to beta decay; and the neutron captures continue until the rate for the (γ, n) reaction balances that for the (n, γ) reaction. The nucleus then waits until β decay again allows it to capture neutrons and the process terminates either when the neutron density and temperature fall low enough for the (n, γ) reactions to cease or when fission occurs when more massive isotopes ($A \sim 275$) have been produced whose spontaneous fission halflife is shorter than the timescale for neutron capture. The neutron capture rate for the r-process can be estimated as follows: the capture is : $\frac{1}{t} = n \langle \sigma V \rangle$, where n is the neutron number density, $\sigma \sim 10^{-25}$ cm^2 is a typical neutron capture cross-section and $V \sim 10^8$ cm s^{-1} for an assumed temperature $T \sim 10^9$ K. Since the neutron capture timescales must be much shorter than beta decay timescales of $t \sim 10^{-5}$ s, which implies $n \simeq 10^{23}$–10^{24} cm^{-3}. Thus very high neutron densities are required and can be produced by only extreme physical conditions like a supernova. The nuclides, that remain after the r-process has ceased, betadecay back to stability in a predictible manner and since nuclides that are close to magic numbers are more likely to remain at the termination of the r-process than nuclides between magic numbers, r-process abundance peaks can be predicted. The observed abundance peaks are in agreement with predictions. The r-process nuclei exhibit abundance peaks near atomic weights 80, 130 and 195 that are correlated with the neutron magic numbers 50, 82, and 126 respectively. Abundance peaks associated with these same neutron magic numbers are also observed for the s-process nuclei, but at somewhat larger atomic weights. This reflects the fact that although the s-process path follows the line of beta stability, the path of the r-process is far to the neutron-rich side of the stability line. In other words, after a large number of neutrons (~ 20) have been rapidly added to an isotope, the rate of neutron decay becomes as rapid as the rate of neutron capture and consequently nuclides remain till further beta decay can occur. After beta decay further neutron capture produces still heavier isotopes and proceeding in this way, the r-process nucleosynthesis goes along a path in the N–Z plane approximately 20–30 neutrons in excess of the valley of beta stability. Beta decay can be neglected during the r-process as beta decay lifetimes $> 10^{-2}$ s. Thus very neutron-rich isotopes are produced in a short time, a nuclide with $A \simeq 250$ requiring about 5s.

For completeness it must be mentioned that it was realized that one could achieve the neutron-rich r-process synthesis without having extremely rapid neutron capture. For instance, Blake & Schramm (1976) showed that nuclei usually attributed to the r-process can be produced by a more general neutron capture process, the n-process, where the neutron capture rates are comparable to the beta decay rates along the synthesis paths. Thus instead of requiring neutron densities $n \sim 10^{24}$ cm^{-3}, we need $n \sim 10^{19}$ cm^{-3}. This would extend the range of possible sites for the r-process but one would still be forced on to some catastrophic event. No single type of site for production of the r-process nuclei has been decided upon. Among the proposed astrophysical sites are:

(i) Neutron rich mass in a supernova explosion (Delano & Cameron 1971). The electron capture rates in the neutronizing core can yield $n/p \gg 1$. An r-process can occur after the protons and some of the neutrons have synthesized seeds. (ii) The neutronized jet from magnetic rotating collapse (Meier *et al.* 1976). (iii) The "tube of toothpast" jet of neutrons in a blackhole-neutron star collision (Lattimer & Schramm 1974). (iv) Supernova shocks traversing the He shell (Blake *et al.* 1979) and causing $^{22}Ne\ (\gamma, n)\ ^{25}Mg$ to produce neutrons, or shocks traversing the C/Ne shell producing neutrons from O^{18} and Ne^{22}. These result in a n-process rather than r-process. As already mentioned, the r-process terminates when eventually, for high-A nuclei, neutron-induced fission becomes faster than beta-decay and even spontaneous fission can be faster than beta decay. The fission products can become seeds for further r-processing and in this way super heavy nuclei (with $N = 184$) can be built up. As the lowest mass of the fission yield curve is below the abundance peak at $A \sim 130$, this means that by fission cycling the r-process could produce the two known top abundance peaks plus the rare earth humps and the actinides. Again the disintegeration of $^{56}Fe \rightarrow 13\alpha + 4n$ can lead to an r-process, for if the temperature drops, the α s-recombine again to form iron group seed nuclei which are now bathed in a sea of α-particles and neutrons, the ratio of neutrons to seed nuclei being large enough for γ-process to occur. Many meteroritic isotopic anomalies may be evidence of the r-process having been at work. Even on earth we have some evidence of natural uranium fission at work in the Oklo uranium mine in Gabon (the 'Oklo' phenomenon; Maurette 1976). Since the mine is 3 billion years old, it had a higher content of U^{235} at that time and underwent fission releasing neutrons and producing depletion of samarium 148, the samples showing isotopic anomalies of samarium isotopes.

We have thus seen that the neutron has played a crucial role in astrophysics producing the whole range of heavier elements of the periodic table.

4. Neutron stars

Another manner in which neutrons have affected astrophysical development is the existence of neutron stars as the endpoint of the evolution of massive stars and the interpretation of pulsars as rapidly rotating neutron stars. A neutron star is a degenerate configuration like a white dwarf except that it consists almost entirely of "cold" degenerate neutrons, all electrons and protons having been converted into neutrons through the inverse beta decay reaction: $p + e^- \rightarrow n + \nu$, and the neutrinos escaping the star. This selfgravitating mass of neutrons is a neutron star. In the white dwarf stage the star (supported by electron Fermi pressure) consists of positive nuclei (helium in the lightest stars, iron in the heaviest) permeated by a relativistic degenerate electron gas throughout the star. If the Fermi energy of the electron gas should exceed the neutron-proton mass difference, inverse beta decay takes place and the electrons are absorbed by the nuclei converting protons into neutrons yielding progressively more and more neutron-rich nuclei; and the process proceeds extremely rapidly because of the very high electron density and energy. Many of the white dwarfs are made up of C^{12} and O^{16} which are stable nuclei and have high threshold energies for inverse beta decays. The critical energy of C^{12} is 13 MeV and is reached at $\rho \approx 10^{11}$ g cm^{-3} and for iron one needs only 3.7 MeV to make inverse beta decay (converting protons in to neutrons) possible. Therefore Fe^{56} will inverse beta decay at only $\sim 10^9$ g cm^{-3}. It may be remarked that the critical density for the inverse beta decay and the critical density for general relativistic effects to become important in white dwarfs are very close to each other, the two effects yielding more or less the same maximal limiting mass. Thus in a simplified model the electron capture process will continue until all the electrons are absorbed and all protons converted into neutrons and such a star would be stabilized by neutron Fermi pressure. That such a contingency could arise in a star was first conjectured by Landau (1932), and a little later Baade & Zwicky (1934) hypothesized that a supernova leaves behind a neutron star which should be found at the centre of the expanding envelope remnant of the supernova explosion. Nearly 34 years after this prophetic paper, a rotating neutron star was found at the centre of the Crab nebula supernova, Gamow (1937) estimated the radii of neutron stars as 10 km and showed that their densities would be of order 10^{14} g cm^{-3}. Oppenheimer & Volkov (1939) realized that general relativity would be important for such dense configurations and obtained the first rigorous mass limits for a neutron star. By anology with white dwarfs, one can write expressions for total energy density and pressure for an ideal Fermi gas of neutrons, with maximum momentum k_F (m_n = neutron rest mass)

$$\rho = \frac{1}{(2\pi\hbar)^3} \int_0^{k_F} (k^2 + m_n^2)^{1/2} \; k^2 dk; \; p = \frac{8\pi}{3} \frac{1}{(2\pi\hbar)^3} \int_0^{k_F} \frac{k^2}{(k^2 + m_n^2)^{1/2}} k^2 dk,$$

By parameterizing in terms of the nuclear density,

$$\rho_c \equiv \frac{8\pi m_n^4 c^3}{3(2\pi\hbar)^3} = 6 \times 10^{15} \text{ g cm}^{-3},$$

we can obtain the equation of state in the form $\frac{\rho}{\rho_c} + F\left(\frac{\rho}{\rho_c}\right)$, with F a definite transcendental function. The structure of a neutron star with a given central density $\rho(0)$ can then be calculated and the mass and radius expressed as functions of $\rho(0)$. As for white dwarfs the problem is analytically solvable only for very large and very small central densities, and an exact solution exists in the limit $\rho(0) \to \infty$. For $\rho(0) \ll \rho_c$ analogy with white dwarfs can be used and M and R turn out to be

$$M = \tfrac{1}{2} \left(\frac{3\pi}{8}\right)^{1/2} (2.714) \left(\frac{\hbar^{3/2}}{m_n^2 \, G^{3/2}}\right) \left(\frac{\rho(0)}{\rho_c}\right)^{1/2}$$

and $$R = (3\pi/8)^{3/2} \, (3.65) \, (\hbar^{1/2}/m_n^2 \, G^{1/2}) \, (\rho_c/\rho(0))^{1/6}.$$

For a pure ideal-gas neutron star, $\longrightarrow M \sim 1 \, M_\odot$, $R \approx 10 \, km$ and stability is determined as usual relative to $(\partial M/\partial \rho(0)$. Of course the neutron star would also contain enough electrons and protons so that the Pauli principle prevents neutron beta decay $n \to p + \bar{e} + \bar{\nu}$, the condition for the neutrons to be stable against such decay being that the electron Fermi momentum $k_{Fe} > k_{max}$. Here k_{max} is the maximum electron momentum in neutron beta decay:

$$k_{\max} \simeq [\Delta^2 - m_e^2]^{1/2} \simeq 1.2 \text{ MeV}.$$

The proton-neutron ratio is large and decreasing for very small neutron densities, reaches a minimum for $m_n n_n$ equal to the transition density $\rho_t \simeq \rho_c [4(\Delta^2 - m_e^2)/m_n^2]^{3/4}$, $\approx 10^{-4}\ \rho_c$, where the ratio of $\left(\frac{n_p}{n_n}\right)$ reaches a minimum given by:

$$(n_p/n_n)_{\min} \simeq [\Delta + \tfrac{1}{2}(\Delta^2 - m_e^2)^{1/2}/m_n]^{3/2} = 0.002\,;$$

(n_p/n_n) then rises monotonically reaching the value (1/8) for $n_n m_n \gg \rho_c$. Thus the 'equilibrium' composition of a neutron star (determined by $d\rho/dn_n = 0$) would consist of 88% neutrons and about 12% protons Many remarkable properties follow from the magnitude of the neutron-proton mass difference. The transition density ρ_t may be pictured as a kind of minimum central density ($\sim 10^{11}$ g cm^{-3}) for neutron stars (drawing the line between extreme high density white dwarfs and neutron stars), the density at which equilibrium shifts to free neutrons. This corresponds to a minimum neutron star mass of $\sim 3M_\odot\ (\rho_T/\rho_c)^{1/2}$ (Harrison *et al.* 1965). Merely considering an ideal gas of neutrons does not seriously affect the overall structure of a neutron star but inclusion of nuclear forces between nucleons are very important especially at densities $\sim \rho_c$ does. The presence of a hard core repulsive potential raises the limiting maximum mass to a model-dependent quantity (usually between 2–3 $M_\odot$) well above the Oppenheimer-Volkov limit. However even the models which consider the details of nucleon interactions are still highly idealized, for a real neutron star is supposed to be quite a complex structure consisting of a crystalline crust (Ruderman 1969), a superfluid interior (Cameron 1970), powerful magnetic fields and very rapid rates of rotation. The maximum rotation frequency and fundamental vibration frequency of any Newtonian polytrope are both of order $(GM/R^3)^{1/2}$ and this result apparently holds to within an order of magnitude for any stable neutron star. For extreme values of M and R, this is $\sim 10^4$ s^{-1}. Pulsars could have started with a rotation frequency near this maximum and subsequently slowed down losing their energy through gravitational and electromagnetic dipole radiation and through radiation by charged particles in the strong magnetic fields. This interpretation is supported by the observation that pulsars are slowing down. White dwarfs with the same mass as a neutron star would have a fundamental vibration frequency and maximum rotation frequency smaller by a factor of 4×10^{-5}; slower than that observed for pulsars.

5. Neutron oscillation and low energy galactic antiprotons

We shall now mention another recent astrophysical application connected with neutrons, *i.e.* a possible relationship between the postulated phenomenon of neutron oscillations and the observation of low energy antiprotons in galactic cosmic rays (Sivaram & Krishan 1982, Sivaram 1983a).

Theoretical predictions of the generation of antiprotons ($\bar{p}$) in cosmic rays as a result of collision of primary protons (p) at relativistic energies would require $\bar{p}/p$ to decrease with decreasing energy, as the threshold for production of $\bar{p}$ from these mechanisms is ~ 2 GeV. This is in contrast with observations where $\bar{p}/p$ increases at lower energies, the ratio being 10^{-4} at 130–320 MeV. The observations of Buffington *et al.* (1981) would imply a $\bar{p}$ energy density of 10^{-4} eV cm^{-3} inside our galaxy. It is well recognized that there is no mechanism to explain this sub-GeV $\bar{p}$ excess, although many suggestions have been made such as secondary production in high energy collisions (Eichler 1982) and blackhole evaporation (Kiraly 1982). Most of these suggestions (including deceleration mechanisms to reduce the primary $\bar{p}$ energy) have inherent difficulties, such as the imposition of severe constraints on the sources to suppress excessive gamma-ray production. Thus one should preferably have a means of producing the $\bar{p}$ s directly at sub-GeV energies without having to invoke decelerating mechanisms. It is obvious that the direct low energy production of $\bar{p}$ s would violate baryon number conservation. Fortunately recent progress in grand unified theories of fundamental forces has provided foı just such a situation (Glashow 1979). The phenomenon of neutron-antineutron $\bar{n}(n\bar{n})$ oscillation has been suggested (Kuzmin 1978; Mahapatra & Marshak 1980) as a possible signature of grand unification in which breakdown of conventionally assumed quantum numbers like baryon number occurs. This transition is a first order process involving a baryon number change of 2. The estimated transition times for the $n\bar{n}$ oscillation are model dependent and range between 10^5–10^7s (corresponing to proton decay time $\sim 10^{31}$ yr). We can write a general Hamiltonian for $n\bar{n}$ mixing in the form:

$$H = \begin{bmatrix} E_0 + \Delta E & \Delta e \\ \Delta e & E_0 - \Delta E \end{bmatrix}, \text{ where } \Delta e = \hbar/\tau_{n-n'} = 10^{-32}\text{–}10^{-34} \text{ erg},$$

for $\tau_{nn'}$ the oscillation time ranging between 10^5–10^7 s, E_0 is the free neutron energy and ΔE is the perturbed energy. Then the intensity of the antineutron component $I(\bar{n}, t)$ in an initially pure neutron beam at $t = 0$ after a propagation time t is given by

$$I(\bar{n}, t) = I(n, 0) \frac{\Delta e^2}{\Delta E^2 + \Delta e^2} \sin^2 (\Delta e^2 + \Delta E^2)^{1/2} t.$$

For $\Delta E \neq 0$, the maximum value of $I(\bar{n}, t)$ is

$$I(\bar{n}, t)_{\max} = (\Delta e/\Delta E)^2.$$

$\Delta E = g\mu_N B$, in an interstellar or intergalactic magnetic field B, μ_N is the nucleon Bohr magneton, g is the neutron anomalous gyromagnetic ratio. If B is in gauss $\Delta E = 9 \times 10^{-24} B$ erg. Thus if we have an intense astrophysical flux of neutrons we should end up with antiprotons (as antineutrons decay in a few minutes to antiprotons). These antiprotons would then have been produced at low energies straight away. We have seen that a supernova explosion releases a very large number of neutrons and one expects a n/p ratio of ~ 10 (Truran 1977). If each supernova explosion produce N neutrons, f be the frequency of explosions, B the interstellar magnetic field and t the diffusion time in the galaxy of the particles, then the number of antiprotons produced is (Sivaram & Krishan 1982; Sivaram 1983a)

$$\bar{p} = N\left(\frac{\Delta e}{g\mu\beta}\right)^2 \times ft.$$

The observed $\bar{p}/p$ ratio requires 10^{55} low energy antiprotons to be present in the galaxy (an observed energy density of 10^{-4} eV cm^{-3}). Putting in the figures, it turns out that the observed $\bar{p}$ background can be built up after a few thousand supernova explosions, *i.e.* over time periods $\sim 10^5$ yr. The same order as that required for the particles to spread over the galaxy.

Conversely, assuming the $\bar{p}$ background to be produced in the aftermath of neutron rich supernova explosions would give a constraint on neutron oscillation times $\sim 10^6$ s, marginally in agreement with the unified theory predictions.

6. Electric dipole moment of the neutron and the entropy per baryon

Ramsey *et al.* (1951) did the first experiment to look for a neutron electric dipole moment (EDM), using the magnetic resonance method. A beam of polarized neutrons is sent into a region with a uniform magnetie field $\mathbf{H}_0$ in such a way that the direction of the polarisation vector of the neutrons is perpendicular to the magnetic field. Simultaneously a uniform electric field is applied parallel to the magnetic field. If there is no EDM, the neutrons precess relative to the direction of the magnetic field with the Larmor frequency. If the neutron has an EDM **d**, the energy for the neutron in the field has an additional term proportional to **d.E** and the precession frequency is changed; this can be observed through a change of the phase of the precession at a certain point of emergence from the field. The present refined experimental situation provides limits of great precision $|d| < 6 \times 10^{-25}$ cm. Landau showed that an EDM of an elementary particle can exist only in the case when there is a violation of not only parity conservation (P) but also time reversal (T). To see this we define the EDM of a particle as follows: Let ρ_J be the charge density inside a particle with angular momentum $\bar{J}$ with quantum number J whose orientation state is given by $m = J$ relative to the z-axis through the centre of mass then $ed = \int \rho_J dvz$; dv is the volume element. If the particle is charged, then this implies that the charge centroid does not coincide with the mass centroid when $d \neq 0$. If the particle is like a neutron uncharged, then $d \neq 0$ corresponds to an excess positive or negative charge in one hemisphere. To connect with CP operation, we see that $d = 0$ if there is symmetry under P or T (d does not change under T but J does, so D must vanish if there is T symmetry). So if CPT is a good symmetry, then T violation implies CP violation and thus $d \neq 0$ if and only if CP is violated. The present experimental limit on d is indeed very accurate, *i.e.* it implies that if the neutron were blown up to the size of the earth, then non-zero charge bulge would correspond to $\sim 10^{-1}$ mm. How does all this tie up with the photon-to-baryon ratio in the universe ? To see this, we must recall that the baryon asymmetry observed in the universe today (we have predominantly matter with $N = 10^{80}$ baryons) is supposed to have been generated in the very early universe through baryon-number-violating interactions mediated by super heavy gauge x-bosons ($M_X \sim 10^{15}$ GeV) which are predicted by the grand unified theories to have existed when the universe was passing through the epoch at $t \approx 10^{-35}$ s (Ellis *et al.* 1981; Weinberg 1979). As pointed out by Sakharov (1967) a necessary condition for obtaining a net baryon number is that CP must be violated (to produce unequal number of charges of opposite sign) in these interactions, thus we must have at least three conditions (i) violation of baryon number, (ii) violation of CP, (iii) thermal non-equilibrium. Decay of the massive X, $\bar{X}$ bosons mediating these interactions leads then to a net baryon number and if ϵ be the mean net baryon number produced in an $X\bar{X}$ decay, we end up generating a specific entropy $S^{-1} \equiv n_B/S \sim 10^{-2}\,\epsilon$ (Weinberg 1979). A CP violation of order $\epsilon \sim 10^{-7}$ is necessary to explain the observed specific entropy of $S \sim 10^{-9}$. We have seen that the existence of a neutron electric dipole moment is tantamount to CP-violation. Then the required amount of CP violation to have produced the necessary baryon number asymmetry for matter (and hence life) to exist sub sequently in the universe is translated in terms of an equivalent electric dipole moment given by (Ellis *et al.* 1981).

$$d \geqslant 2 \times 10^{-18}\,(n_B/S)\text{ cm.}$$

Direct observations and the result of cosmological nucleosynthesis imply $n_B/S \geqslant 3 \times 10^{-10}$, and hence a predicted $d \geqslant 6 \times 10^{-28}$ cm. This condition on d is much more stringent than present experimental bounds. Conversely existing experimental upper limits on d place a lower limit on the specific entropy generated by GUTS at $\sim 10^{-35}$ s of $S \leqslant 3 \times 10^{6}$. If future experiment improve the limit on the neutron EDM to beyond $< 5 \times 10^{-28}$ cm; and the neutron is seen to have no EDM in this limit, we will have to seriously consider the possibility of GUTS being able to generate the observed baryon asymmetry and the photon-to-baryon ratio seen in the universe. A more detailed aspect of this is under preparation. Another interesting consequence of the sensitivity of the neutron oscillation phenomenon to magnetic fields is that it enables us to set a lower limit on any primordial magnetic field present in the neutron-rich nucleosynthetic era of the big bang on the basis that it should not interfere with the helium production. This limit is obtained as an interesting combination of fundamental constants : (Sivaram 1983c),

$$B \geqslant \left(\frac{m_n}{m_\pi}\right) \frac{(480\pi^3 m_p^{\frac{1}{2}})^{\frac{1}{2}}}{g_\tau e G_F K T}\, \hbar^3 c^3 \left(\frac{G}{\hbar c}\right)^{1/8}.$$

Assuming flux conservation, this extrapolates to a present value of $B_0 = B\,(T_0/T)^2$ and for a present day background temperature, $T_0 = 3$ K, $B_0 > 10^{-10}$ gauss for $\tau > 10^6$ seconds.

It may also be mentioned that the DUMAND project is attempting to detect relativistic neutrons in the very high energy range from supernovae and that the 2.25 MeV gamma ray line from the radiative capture of neutrons by protons has already been seen, *e.g.* in solar flares.

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