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An Introduction to Nuclear Astrophysics

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AN INTRODUCTION TO NUCLEAR ASTROPHYSICS*

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ABSTRACT

The role of nuclear reactions in astrophysics is described. Stellar energy generation and heavy element nucleosynthesis is explained in terms of specific sequences of charged-particle and neutron induced reactions. The evolution and final states of stars are examined.

1. INTRODUCTION

In this series of lectures, I will attempt to provide an introduction to the field of nuclear astrophysics. I will start with a review of the astronomical observations that provide the basis for the subject. Then the sequence of static nuclear burning stages of stars will be examined. The explosive deaths of massive stars, known as supernovae, will be described. The final states of stars will then be discussed. We will see that a consequence of this sort of stellar evolution is the production of most of the elements between carbon and iron. However, in order to explain the observed abundances of still heavier elements, it is necessary in invoke the idea of neutron-capture reactions. In fact, two distinct types of neutron capture processes seem to be needed: the so-called s- or slow process and the r- or rapid process. There are several excellent books on the subject of nuclear astrophysics[1,2,3] and I have relied heavily on them as references in preparing these lectures.

2. ASTRONOMICAL BACKGROUND

The first thing that one notices when looking up at the heavens is that the sky is dark at night. While this may at first seem trivial, the more deeply one examines this fact the more puzzling it actually becomes. This problem has a long history, and has come to be known as Olber's Paradox. While I won't have time to discuss this subject, I would point out that most textbook answers to this problem are wrong or at least misleading. For an excellent discussion of this fascinating topic I would refer you to the work of E. Harrison.[4)]

The next thing we notice are the little points of light in the sky which we call stars. By examining the light that comes to us from the stars, it is possible to deduce a great deal of information about the nature of the universe. Remarkably, it seems that the same basic laws of physics apply everywhere we look. Furthermore, all the objects that we can observe seem to be made up of the same 92 chemical elements found on Earth. From such observations we now know that approximately 73% of the mass of the visible universe is in the form of hydrogen. Helium makes up about 25% of the mass, and everything else represents only 2% of the mass of the universe. While the abundance of these "heavy" elements seems quite low, it is important to remember that most of the atoms in our bodies and in the Earth are a part of this small portion of the matter in the universe. As we will see later, it is now believed that the hydrogen and helium were produced in the Big Bang and that the heavy elements are the results of nuclear reactions that take place in stars.

While nuclear astrophysics is now a well established subfield of nuclear science, until fairly recently its foundation was based almost entirely on circumstantial evidence. The reason for this is quite simple. We cannot look inside stars to see what is actually going on. We must rely on more indirect means. One of the strongest pieces of evidence that nucleosynthesis does occur in stars is the observation of the element technetium on the surfaces of certain old stars.[5)] Technetium is one of only two elements below bismuth that has no stable isotopes. In fact the longest lived technetium isotope has a half-life of only 4×10^6 years.

While this may seem long by human standards, it is quite short on astronomical timescales. The only plausible way for such "short-lived" material to be present in a star is for it to have been recently synthesized within that star. More recently, observations of a 1.809-MeV gamma ray coming from the galactic plane show that the short-lived ($t_{1/2} = 7.2 \times 10^5$ years) isotope ^{26}Al is present.[6] This indicates that nucleosynthsis is still occurring in our galaxy.

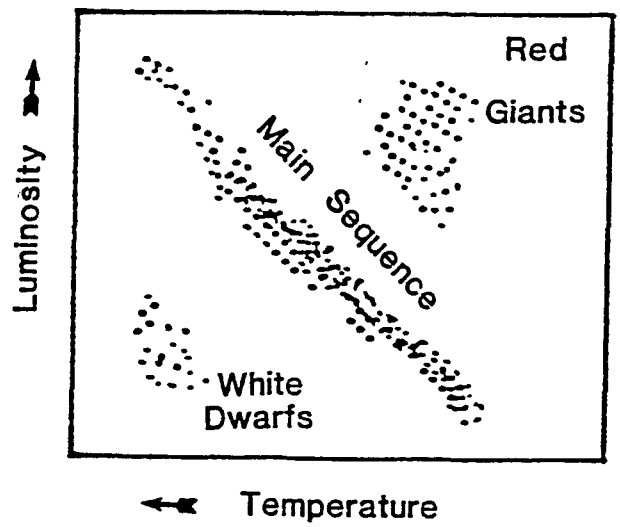

Figure 1. The Hertzsprung-Russell diagram.

Another important piece of astronomical data is the observed relationship between the surface temperatures and luminosities of stars. Shown in Figure 1 is what is known as a Hertzsprung-Russell diagram. It has been found that about 80% of all the observed stars fall on a roughly diagonal band known as the main sequence. More rare are two classes of stars known as red giants and white dwarfs. We will come back to this diagram a number of times during these lectures.

Finally, a large amount of information can be learned from more detailed analysis of the elemental and isotopic composition of matter. Shown in Figure 2 are the observed abundances of the material in our solar system. As discussed previously, hydrogen and helium are by far the most abundant species. The next heaviest group of elements, Li, Be, and B, are by comparison exceedingly rare. Above this group, the abundances start out higher but gradually decrease as one moves up to heavier nuclei. Several

important features of this abundance pattern should be pointed out however. There is a large abundance peak near mass 60. Above this point, there is again a general decrease in abundance with increasing mass number that is interrupted by two double-peaked structures. Distinct abundance peaks are observed around mass numbers 130, 140, 195, and 208. Taken as a whole, this abundance distribution provides many clues to the source of stellar energies and the origin of the chemical elements.

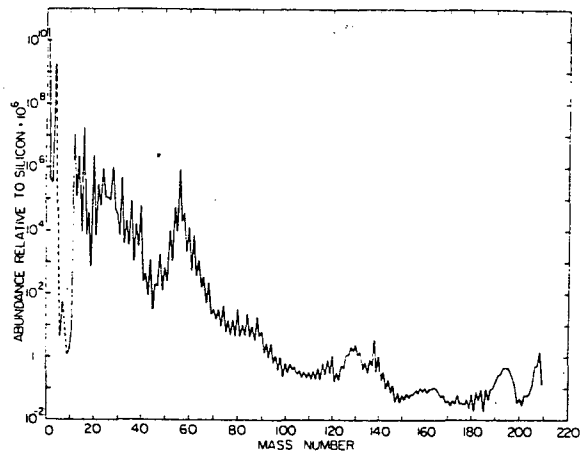

Figure 2. Solar system abundances (by number) of the nuclides [Ref. 7].

3. WHAT MAKES THE STARS SHINE?

One of the most basic questions that we can ask about stars is what is the source of their energies, or in other words, what makes the stars shine? Before one can answer this question, some basic properties of stars must be known. Taking our Sun as a representative star, we know that:

$$M_{\odot} = 2 \times 10^{33} \text{ grams} \quad (1)$$

$$L_{\odot} = 4 \times 10^{33} \text{ ergs/second} \quad (2)$$

$$\text{Age}_{\odot} = 4.5 \times 10^{9} \text{ years} \quad (3)$$

We know that life has existed on Earth for at least the last 2×10^9 years and this requires that the Sun's luminosity not have changed dramatically over that period of time. Therefore, over the lifetime of the Sun, it has

radiated a total of 6×10^{50} ergs or 3×10^{17} ergs/gram.

There are several possible sources of this energy that must be considered. Exothermic chemical reactions are responsible for much of the energy generation on Earth. However, the maximum energy release in such reactions is approximately 2×10^{12} ergs/gram. Thus, chemical reactions could keep the Sun at its present luminosity for only about 30,000 years. Thus, chemical reactions cannot be the source of stellar energies.

Gravitation is another possible energy source. If one assumes that the matter in the Sun contracted from infinite initial separation down to the present radius, then the change in the gravitational potential energy could be the source of the Sun's energy. However, one can easily calculate that the total gravitational potential energy of the Sun (assuming constant density) is

$$V = 3/5(GM^2/R) = 2 \times 10^{48} \text{ ergs} \tag{4}$$

This translates to about 10^{15} ergs/gram, still far short of the needed amount.

Near the beginning of this century it was discovered that nuclear reactions are capable of producing large amounts of energy, and the possible role of nuclear reactions in stars was soon realized. Consider the possibilty of combining four hydrogen nuclei in such a way as to make a nucleus of ^{4}He. We will see later on that this does not actually occur in one step, but in fact requires a number of separate reactions. However, for the present calculation, all that matters is the fact that for each ^{4}He nucleus produced in this way, E = 26 MeV is released. One can now calculate the energy generation efficiency of this process,

$$E/4m_pc^2 = 7 \times 10^{-3} \tag{5}$$

If one could convert 1 gram of matter completely into energy, 9×10^{20} ergs would be released. Thus, the fusion of hydrogen into helium yields

$$(7 \times 10^{-3})\ (9 \times 10^{20}\ \text{erg/g}) = 6.3 \times 10^{18}\ \text{erg/g} \qquad (6)$$

Therefore, only about 5% of the hydrogen in the Sun need be burned into helium in order to meet the energy generation requirement. Thus, it has been concluded that the sources of stellar energies are nuclear reactions. We shall also see that the "ashes" of these reactions are the elements between carbon and iron.

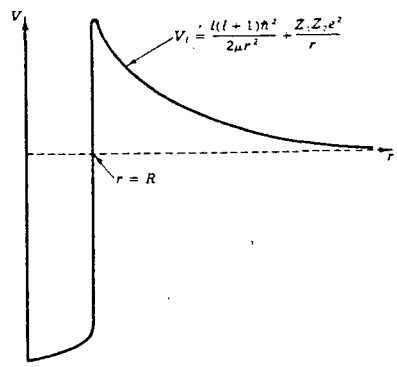

Figure 3. The potential between two nuclei as a function of their separation [Ref. 1, p. 322].

Figure 3 illustrates the potential of two nuclei as a function of their separation. At large distances, they repel one another via the long range Coulomb force, while at short distances the strong, attractive nuclear force takes over. In order for for a nuclear reaction to occur, the two nuclei must reach a separation approximately equal to the sum of their radii. The energy required to bring two nuclei of charges Z_1 and Z_2 and masses A_1 and A_2 to this point can be easily calculated, and is known as the Coulomb barrier:

$$E_c \approx (Z_1 Z_2)/(A_1^{1/3} + A_2^{1/3})\ \text{MeV} \qquad (7)$$

As an example, consider the interaction of two protons , which we will see is the first step in the synthesis of helium from hydrogen. The Coulomb barrier for this reaction is about 0.5 MeV. This must be compared to the typical thermal energies found in stars. At the center of the Sun,

$$T_{\odot} = 1.5 \times 10^7 \text{ K} \tag{8}$$

Thus, the mean thermal kinetic energy of a nucleus at the center of the Sun is

$$kT_{\odot} = 2 \text{ keV} \tag{9}$$

Under such conditions, classical physics says that the protons can never get close enough to one another for a reaction to occur. Nevertheless, the stars shine! The nuclear reactions that power the stars proceed via quantum mechanical tunneling through the potential barrier. I would point out that this is a general feature of all of the major nuclear burning stages of stars. During helium burning, which occurs at a temperature of about one hundred million degrees, the mean thermal kinetic energy is about 13 keV. Even at a temperature of one billion degrees, which is appropriate for oxygen burning, this thermal energy is only 130 keV.

The stellar reaction rate for a particular nuclear reaction between species 1 and 2 is given by the expression

$$R_{12} = [1/(1 + \delta_{12})]\, N_1 N_2 \langle \sigma v \rangle \tag{10}$$

where δ_{12} = Kronecker delta, N_1 and N_2 are the number densities of species 1 and 2, respectively, and $\langle \sigma v \rangle$ = cross section times velocity averaged over a Maxwell Boltzmann velocity distribution. As can be seen in Figure 4, the cross sections for charged-particle induced reactions increase rapidly with energy due to the Coulomb penetration factor. On the other hand, the Maxwell-Boltzmann velocity distribution peaks at low values of the energy. These two competing factors in the above equation result in the bulk of the reaction rate occurring at an energy $E_o \pm \Delta/2$ that is generally much greater than kT.

$$E_o = 1.22\, (Z_1^2 Z_2^2 A T_6^2)^{1/3} \text{ keV} \tag{11}$$

$$\Delta = 0.75\, (Z_1^2 Z_2^2 A T_6^5)^{1/6} \text{ keV} \tag{12}$$

where A is the reduced mass of the system in atomic mass units, and T_6 is the temperature in units of 10^6 K.

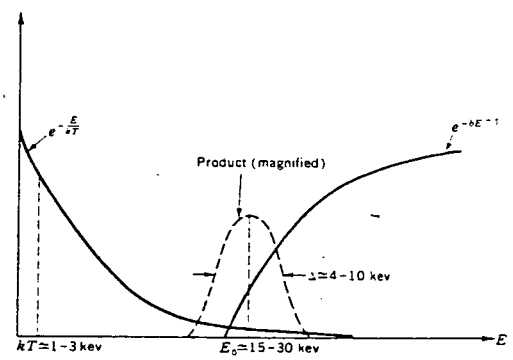

Figure 4. The competition between the Maxwell-Boltzmann energy distribution and the barrier penetration probability determines stellar reaction rates [Ref. 1, p. 302].

Figure 5a illustrates the excitation function for one of the reactions important for hydrogen burning in massive stars, the $^{12}C(p,\gamma)^{13}N$ reaction. A resonance is seen at a laboratory proton energy of 480 keV, corresponding to the $J^{\pi} = 1/2^+$ level at 2367 keV excitation energy in ^{13}N. Below this point, the cross section falls dramatically due to the Coulomb barrier. The astrophysically interesting energy range for this reaction is 25 ± 15 keV. The cross section drops by a factor of about a million in going from 400 to 100 keV, thus making the required extrapolation questionable at best. Since it is known that the reason for this decrease is the Coulomb penetration factor, this term can be factored out to obtain the astrophysical S-factor

$$S = \sigma(E)E \exp(2\pi Z_1 Z_2 e^2 / \hbar v) \qquad (13)$$

Shown in Figure 5b is this S-factor determined for the same reaction. While the cross section changes by a factor of 10^6 in going from 400 to 100 keV, S(E) changes by only ~ 25, thus making the extrapolation to 25 keV more reliable.

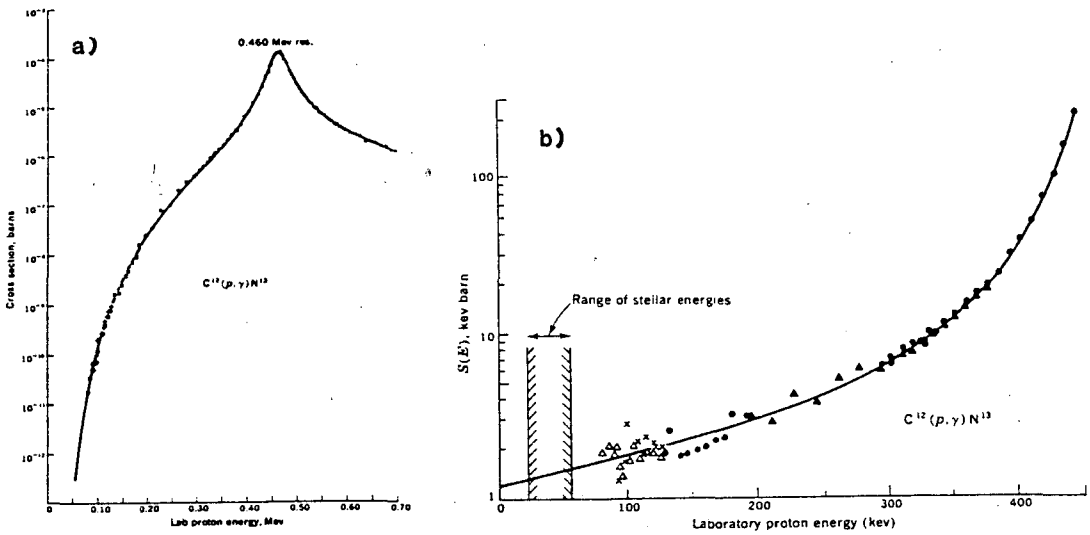

Figure 5. a) The measured cross sections for the $^{12}C(p,\gamma)^{13}N$ reaction. b) The cross-section factor, S(E), for the same reaction [Ref. 1, p. 298,300].

Now let's examine more closely the sequence of reactions that are believed to be going on in our Sun by which hydrogen is converted into helium. The energy generated by the fusion of hydrogen into helium provides main sequence stars their support against gravitational contraction. Shown below are the reactions, the fraction of the time each occurs in the Sun, and the energy released in the reaction (not including the energy carried off by neutrinos). The first reaction involves a weak interaction and is the rate determining step in the sequence. As mentioned above, we cannot "see" what is going on at the center of the Sun. A photon emitted at the center of the Sun scatters many many times as it works its way outward. In fact, such a photon takes $\sim 10^7$ years to go from the center of the Sun to the surface. However, neutrinos do freely flow out of the Sun, and if we could detect them, we could learn much about what really goes on inside a star.

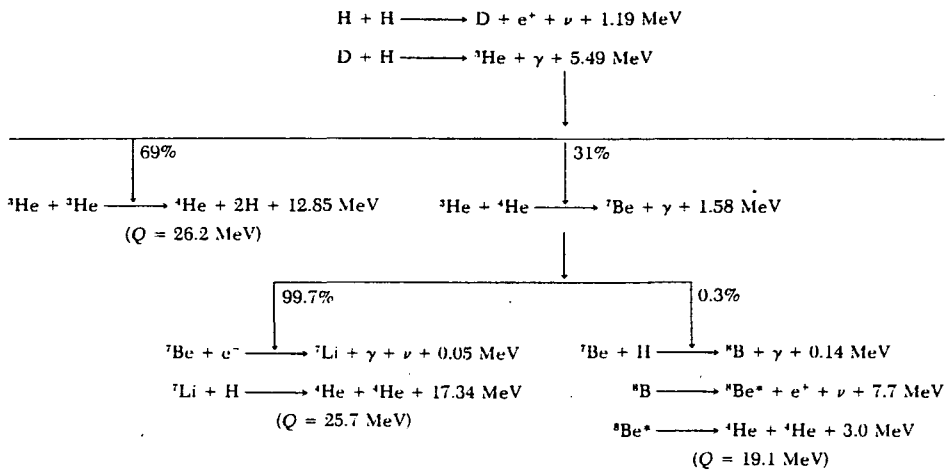

Table 1. The sequences of reactions by which hydrogen is fused into helium in the Sun [Ref. 8].

One can make a simple estimate of the expected flux of solar neutrinos at the surface of the Earth. The only assumption that one has to make is that the observed present luminosity of the Sun is due to the fusion of hydrogen into helium.

$$\Phi_{\nu pp} = 2\, L_{\odot} / Q \tag{14}$$

where $L_{\odot}$ = 1 kilowatt/ m^2 (the power of sunlight at the Earth's surface) and Q = 26 MeV (the energy released in the fusion of four hydrogen atoms into one helium atom). One obtains

$$\Phi_{\nu pp} = 5 \times 10^{10} \ / \ cm^2 \text{ - sec} \tag{15}$$

At present, there is only one operating solar neutrino detector - that of Raymond Davis and his collaborators.[9] It is a radiochemical experiment based on the reaction

$$^{37}Cl + \nu_e \text{ -- } ^{37}Ar + e^- \qquad Q = -\ 0.814 \text{ MeV} \tag{16}$$

The target consists of a tank containing 100,000 gallons of C_2Cl_4 (a dry-

cleaning fluid) located about one mile underground in a gold mine in Lead, South Dakota. The ^{37}Ar produced via ν captures is flushed from the tank about once a month and its decay is measured in a remote counting laboratory. The average measured ^{37}Ar production rate is approximately 0.5 atoms/day! This can be expressed as

$$\Phi_{\nu expt} = 2.0 \pm 0.3 \text{ SNU} \tag{17}$$

where 1 SNU = 10^{-36} neutrino captures/target atom/second. This rate has been estimated many times by different groups of physicists and the result is[9)]

$$\Phi_{\nu theor} = 7.6 \pm 2.5 \text{ SNU} \tag{18}$$

This discrepancy is the famous solar neutrino problem. It should be pointed out that because of the relatively high threshold energy for the ^{37}Cl detector, the Davis experiment is completely insensitive to the pp neutrinos, but instead is most sensitive to the rare but high energy neutrinos produced in the decay of ^{8}B.

A number of possible explanations to the solar neutrino problem have been proposed. It could be that some of the details of the solar models used to calculate the neutrino flux are incorrect (e.g. the heavy element abundances at the center of the Sun might be different than those observed on the surface, etc.). It might also be that the cross section values used to calculate the nuclear reaction rates are incorrect. Finally, a more exciting possibility is that on their way from the center of the Sun to the Earth, neutrinos oscillate from one type (flavor) to another. Since the ^{37}Cl detector is sensitive only to electron-type neutrinos, this could explain the low measured rate.

There are a number of other potential solar neutrino detectors. In particular, ^{71}Ga is attractive because of its low 0.236 MeV threshold energy. Calculations indicate that a ^{71}Ga detector would show a rate of approximately 100 SNU. However, this would still require 50 tons of gallium in order to get one neutrino capture per day. This important

experiment is now being pursued by two different multi-national groups and hopefully in the next few years the results of these experiments will allow us to decide what is really going on with the neutrinos from the Sun.

In stars more massive than our Sun, the fusion of hydrogen into helium takes place through a sequence of reactions in which ^{12}C serves as a catalyst.[10)]

$$^{12}C + p \rightarrow {}^{13}N + \gamma \tag{19}$$
$$^{13}N \rightarrow {}^{13}C + e^{+} + \nu_e \tag{20}$$
$$^{13}C + p \rightarrow {}^{14}N + \gamma \tag{21}$$
$$^{14}N + p \rightarrow {}^{15}O + \gamma \tag{22}$$
$$^{15}O \rightarrow {}^{15}N + e^{+} + \nu_e \tag{23}$$
$$^{15}N + p \rightarrow {}^{12}C + {}^{4}He \tag{24}$$

Through either the pp chain or the CNO cycle, a star will burn hydrogen into helium until the hydrogen in the core of the star is exhausted. At this point the star begins to move off the main sequence and becomes a red giant. Because the energy generation mechanism is turned off, the core of the stars gravitationally contracts. The central temperature rises until helium is "ignited." Since there are no particle-stable A = 5 or 8 nuclei, helium burning cannot proceed via two-body reactions like those described above. From the observed abundance curve, it is known that following hydrogen and helium, carbon and oxygen are the two most abundant elements in the universe. The question is, how are they produced.

The answer was provided by E. Salpeter[11)] and F. Hoyle[12)] who realized that in order to bridge the gaps at A = 5 and 8, a "three-body" reaction was necessary. As can be seen from Figure 6a, a ^{8}Be nucleus is only 92 keV unbound with respect to decay into two ^{4}He nuclei and has a short but finite lifetime of 2×10^{-16} seconds. This means that at the high temperatures and densities required for helium burning, an equilibrium will be established between ^{4}He and ^{8}Be. In fact at $T = 1 \times 10^{8}$ K and $p = 1 \times 10^{5}$ grams/ cm^{3}

$$[^{8}Be] \,/\, [^{4}He] = 10^{-9} \tag{25}$$

Thus, the unstable ^{8}Be could capture another ^{4}He to produce ^{12}C.

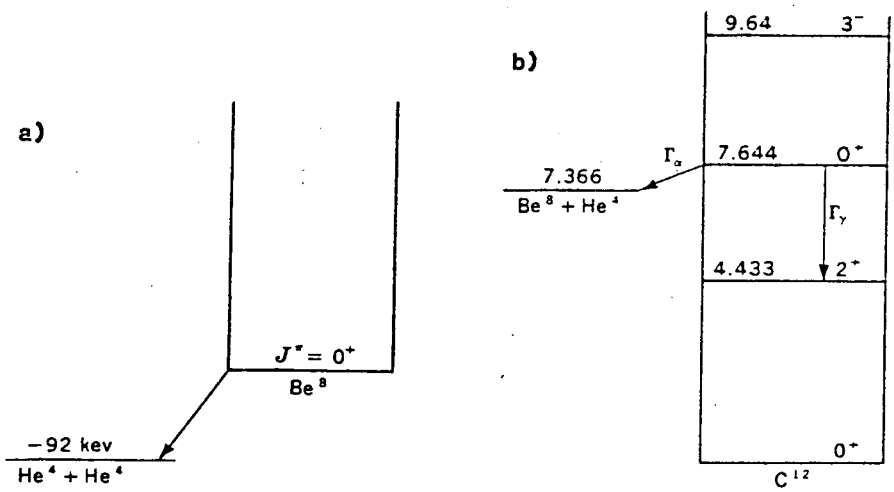

Figure 6. a) The energy level diagram of ^{8}Be. b) The energy level diagram of ^{12}C [Ref. 1, p. 412,413].

It was soon realized that the rate of this process would be unacceptably low unless there were a suitable state in ^{12}C that could serve as a resonance for the ^{8}Be + ^{4}He reaction. To be such a resonance, the state must be a $J^{\pi} = 0^{+}$ level, lie near the ^{8}Be + ^{4}He threshold energy, and have a reasonable gamma-decay branch to the ground state of ^{12}C. The existence of such a level was proposed by F. Hoyle in 1954, and as can be seen in Figure 6b, subsequent experiments have demonstrated that a level does exist in ^{12}C with just those properties required to make the "triple alpha process" work.

Once ^{12}C is formed, ^{16}O can be produced via the $^{12}C(\,^{4}He, \gamma\,)$ reaction. The low-energy cross sections for this reaction were measured by Dyer and Barnes[13] and more recently by Kettner et al.[14] The results of these two investigations disagree in the extrapolated value of the S-factor and further experiments are now underway to resolve the controversy over this important reaction rate.

Helium burning proceeds in the stellar core until all of the helium is converted into ^{12}C and/or ^{16}O. After this point is reached, there are two possibilities for the star. If it is sufficiently massive ($M \gtrsim 10\ M_{\odot}$) the

core will again contract, the temperature will rise, and when the central temperature reaches about 5×10^8 K, the ^{12}C will ignite.

For low mass stars, on the other hand, the core temperature never gets high enough to burn carbon. Therefore, no further energy-generating reactions are possible. Such stars quietly end their lives as white dwarfs with their support against gravitational collapse provided by electron degeneracy pressure. It has been found that there is an upper limit to the mass of a star that can be supported in this way of 1.4 $M_\odot$ that is known as the Chandrashekar limit.[15] More massive stars cannot end their lives in this way and are destined to become either neutron stars or black holes.

For the more massive stars, once the core temperature reaches 5×10^8 K, the $^{12}C + {}^{12}C$ reaction produces large amounts of ^{20}Ne and ^{24}Mg. At 1×10^9 K, oxygen burning begins in which $^{16}O + {}^{16}O$ reactions produce ^{28}Si and ^{32}S. The synthesis of still heavier elements does not proceed directly through the reactions such as $^{28}Si + {}^{28}Si$, because in order to overcome the high Coulomb barriers, temperatures on the order of $4.5 \times {}^{9}K$ are required. At such high temperaturs, photodisintegration reactions become important which allows the following sort of rearrangement reactions to occur.

$$^{28}Si + \gamma \rightarrow {}^{24}Mg + {}^{4}He \qquad (26)$$

$$^{28}Si + {}^{4}He \rightarrow {}^{32}S + \gamma \qquad (27)$$

Under these conditions, nuclear statistical equilibrium can occur which leads to the synthesis of large amounts of the most tightly bound nuclei - the "iron peak" elements.

The time that a massive star spends in each of these static burning stages gets progressively shorter as the star evolves. Shown below are the results of calculations of the evolution of a 25 $M_\odot$ star. While such a star spends millions of years in its initial hydrogen burning stage, it spends only about a day in its final static silicon burning.

Burning Stage	Temperature	Density	Timescale
Hydrogen	5 keV	5 g/cc	7×10^6 years
Helium	20 keV	700 g/cc	5×10^5 years
Carbon	80 keV	2×10^5 g/cc	600 years
Neon	150 keV	4×10^6 g/cc	1 year
Oxygen	200 keV	10^7 g/cc	6 months
Silicon	350 keV	3×10^7 g/cc	1 day
Collapse	600 keV	3×10^9 g/cc	seconds
Bounce	3 MeV	10^{14} g/cc	milliseconds
Explosive	100–600 keV	varies	.1–10 seconds

Table 2. The major stages in the evolution of a 25 $M_\odot$ star [Ref. 16].

Once the core of the star is converted into iron-group nuclei, the star has nearly reached the end of its life. Because the binding energy per nucleon reaches a maximum at this point, there are no further energy generating reactions possible. Thus, once again the core of the star will start to contract and heat up. Eventually the point of iron photodisintegration is reached. This energy drain further removes support against gravitational collapse. The details of what happens next are still not completely clear, but we know what the final result is - a supernova explosion. If one could look inside such a star just prior to the explosion, it is believed that it would exhibit the "onion-skin" structure illustrated in Figure 7. The deeper inside the star one looks, the higher is the peak temperature and correspondingly the heavier are the nuclei that are synthesized.

As the collapse of the core occurs, and the density grows, the Fermi level of the electrons becomes high enough for it to become energetically favorable for electron capture to occur. This neutronizes the core and produces a large flux of neutrinos. Eventually, the core reaches and then surpasses the density of nuclear matter. At this point the equation of state of the matter stiffens, and a hydrodynamic bounce occurs. Through the scattering of the neutrinos or the bounce, or some combination of the two, a supernova occurs in which the mantle of the star is blown off leaving behind a neutronized remnant. If the mass of the remnant is less than 2-3

$M_{\odot}$, it will settle down as a neutron star supported against further collapse by the pressure of degenerate neutrons. More massive stars will continue to collapse to form black holes.

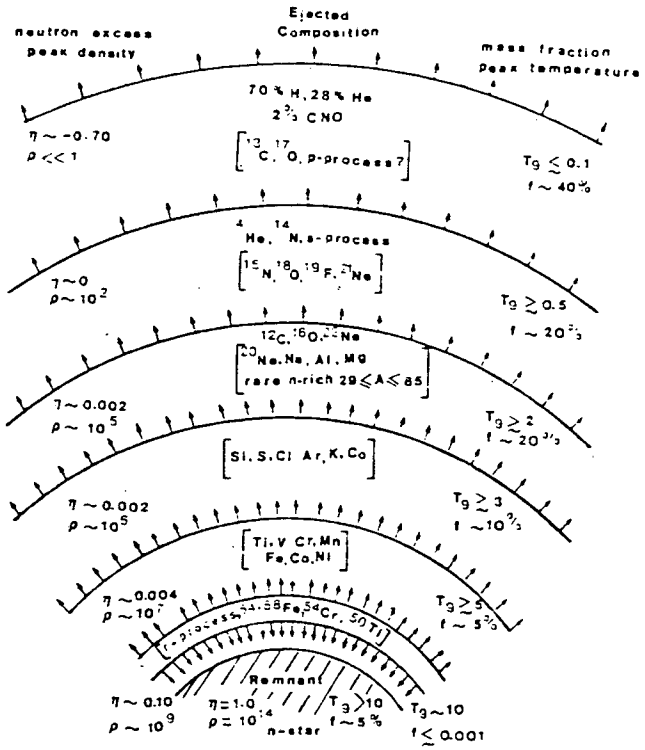

Figure 7. Cross sectional view of a highly evolved star just before the core collapses [Ref. 2, p. 72].

The ashes of these sequences of first static and then explosive nuclear reactions are the bulk of the elements between carbon and iron. While many details remain to be worked out, the overall agreement between the observed abundances and those calculated using large nucleosynthesis networks and the most up to date estimates of the relevant cross sections is reasonably good.

4. ORIGIN OF THE HEAVY ELEMENTS (A > 56)

The type of nuclear reactions discussed so far terminate at iron. As discussed above, because the binding energy per nucleon reaches a maximum at around iron, further fusion reactions between heavy nuclei are endothermic. Furthermore, the Coulomb barriers for charged-particle

induced reactions become prohibitively high. It was realized thirty years ago that in order to account for the observed abundances of the elements above iron, neutron capture reactions are required. The two double-humped peaks seen in the Solar System abundance distribution at A = 130, 140 and A = 195, 208 appear to be correlated with the neutron magic numbers 82 and 126. Based upon this observation, it was suggested by Burbidge, Burbidge, Fowler, and Hoyle[17] that two distinct types of neutron capture processes are required. In the s- or slow process, neutron captures through the isotopes of a given element proceed until a radioactive nucleus is reached. Then, because the neutron flux is so low, beta decay almost always occurs before the next neutron comes along. Thus the path of the s-process follows the line of beta stability. In contrast, during the r- or rapid process the neutron flux is so high that many many neutron captures occur before beta decay. The path of the r-process thus lies far to the neutron-rich side of beta stability. Once the r-process neutron source is turned off, these neutron-rich nuclei beta decay back to the stable nuclei.

Figure 8 illustrates how the s- and r-processes contribute to the observed abundances of the stable isotopes. Most nuclei can be reached through both processes. However, some nuclei are shielded from the r-process by the presence of a stable more neutron-rich isobar. Also, some neutron-rich nuclei cannot be reached in the s-process because of the presence of a neighboring short-lived isotope. There is also a group of generally low-abundance nuclei that cannot be produced in either the s- or r-processes. The synthesis of these nuclei is attributed to the so-called p-process.

During the s-process, a nucleus (Z,N) is produced through neutron captures on the isotope (Z,N-1) and is destroyed through neutron captures on itself. The rate of change of the abundance of specie i is

$$dN_i/dt = n_n N_{i-1} \sigma_{i-1} v - n_n N_i \sigma_i v \qquad (28)$$

where n_n = neutron density, N_i = abundance of specie i, σ_i = (n,γ) cross section, and v = neutron "thermal" velocity.

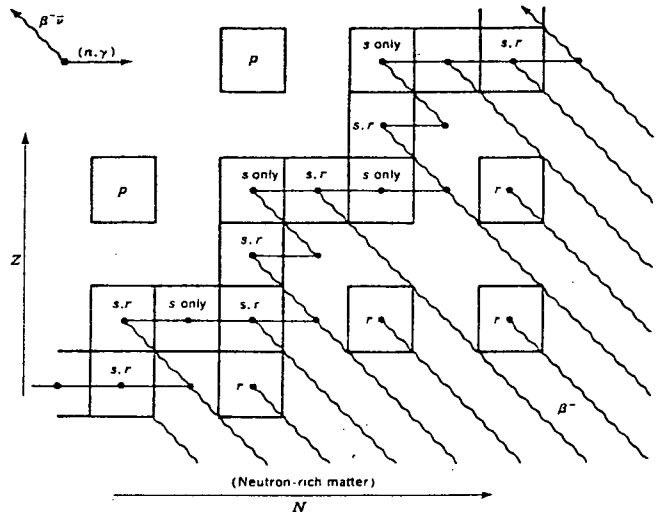

Figure 8. A portion of the chart of the nuclides showing the assignments of nuclei to the s,r, and p processes [Ref. 1, p. 547].

In the limit of a steady flow, $dN_i/dt = 0$. Thus, if such a situation occurred in the s-process then one would expect

$$N_{i-1}\sigma_{i-1} = N_i\sigma_i \tag{29}$$

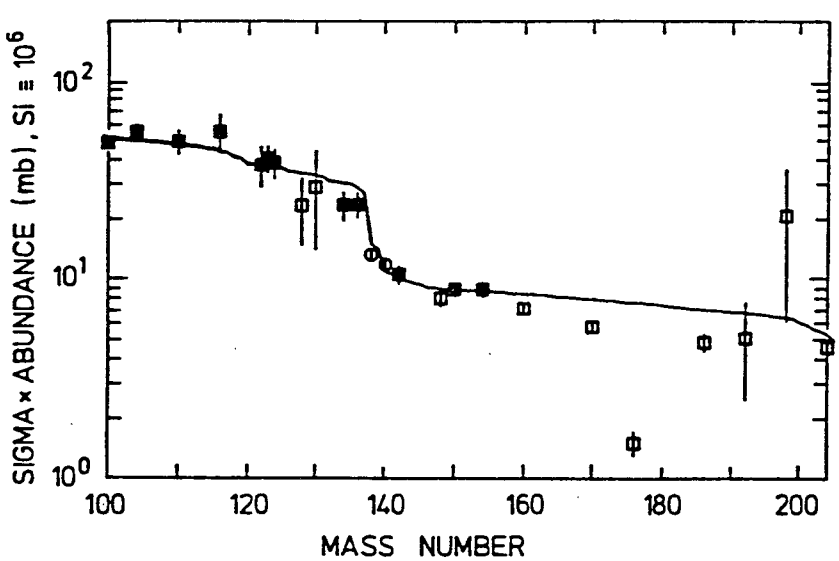

Figure 9. σN versus A for solar system material [Ref. 18].

As can be seen from Figure 9, in between the major neutron shell closures, this is approximately what is observed. For nuclei with neutron magic numbers 50, 82, and 126, experiments have shown that their neutron capture cross sections are very small. Thus, from equation (28) we can understand why abundance peaks are observed for these nuclei. This inverse relationship between abundance and neutron capture cross section is not just a solar system phenomenon, but is also seen in the σN curves that have been obtained for other stars.[19)]

The site of the s-process is belived to be in the helium-burning zones of red giant stars. A cross sectional view of such a star is shown in Figure 10. The neutrons required for the s-process are thought to be produced through (^{4}He,n) reactions on ^{13}C and/or ^{22}Ne, which are produced by the reactions

$$^{12}C(p,\gamma)^{13}N \xrightarrow{\beta^+} {}^{13}C \tag{30}$$

$$^{14}N(^4He,\gamma)^{18}F \xrightarrow{\beta^+} {}^{18}O(^4He,\gamma)^{22}Ne \tag{31}$$

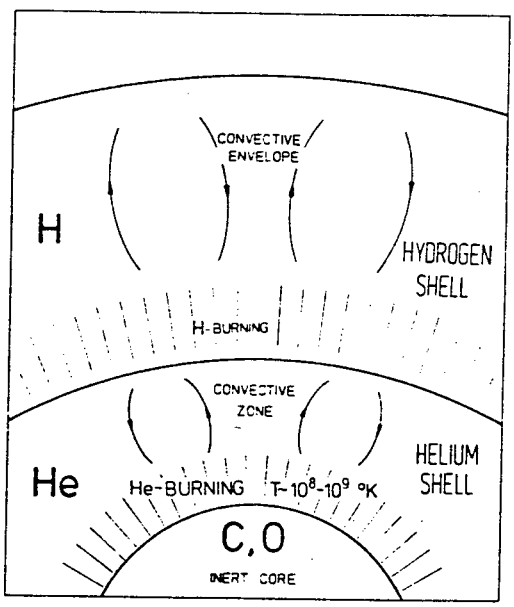

Figure 10. Cross sectional view of a 7 $M_\odot$ red giant star [Ref. 18].

The neutrons generated by these reactions are then captured on "seed" nuclei. Calculations show that it is not possible to reproduce the observed s-process abundances under reasonable helium-burning conditions if these seed nuclei are too light. The most likely seeds are, in fact, iron group nuclei. It is important to remember, however, that no iron is produced in stars during hydrogen or helium burning. Thus, in order for an s-process to occur in a red giant star, these iron nuclei had to be present when the star formed. This could only be the case if the matter from which this star formed had already been "contaminated" by the ashes of previous generations of stars. Thus, the s-process is a secondary nucleosynthetic process and could not have occurred in the first stars that formed in our galaxy.

The s-process terminates at ^{209}Bi because the addition of a neutron to this nucleus produces ^{210}Bi which, through alpha and beta decays, eventually leads back to ^{206}Pb. In order to account for the observed abundances of thorium and uranium, a way of avoiding this point of alpha instability is required. The s-process produces approximately one-half of the nuclei between iron and bismuth. In order to explain the other half of the observed abundances and to understand the origin of the actinide elements, the r-process is required. While we know that a rapid neutron-capture process occurs in Nature, our understanding of this type of nucleosynthesis is much more limited than that for the s-process.

Because the s-process is fairly well understood, one can calculate its contribution to the abundance of each isotope, subtract that from the observed abundance, and attribute the remainder to the r-process. After this is done, clear r-process abundance peaks are seen at A = 80, 130, and 195. However, if one plots σN_r versus A, as was done for the s-process, one sees an essentially random distribution. This indicates that the observed abundances of the stable nuclei that we attribute to the r-process are not correlated with the neutron-capture cross sections of these nuclei. This is a reflection of the fact that nuclei directly produced during the r-process are short-lived neutron-rich nuclei that subsequently beta decay to produce the stable nuclei which we observe. A calculated path of the r-process through the chart of the nuclides is shown in Figure 11.

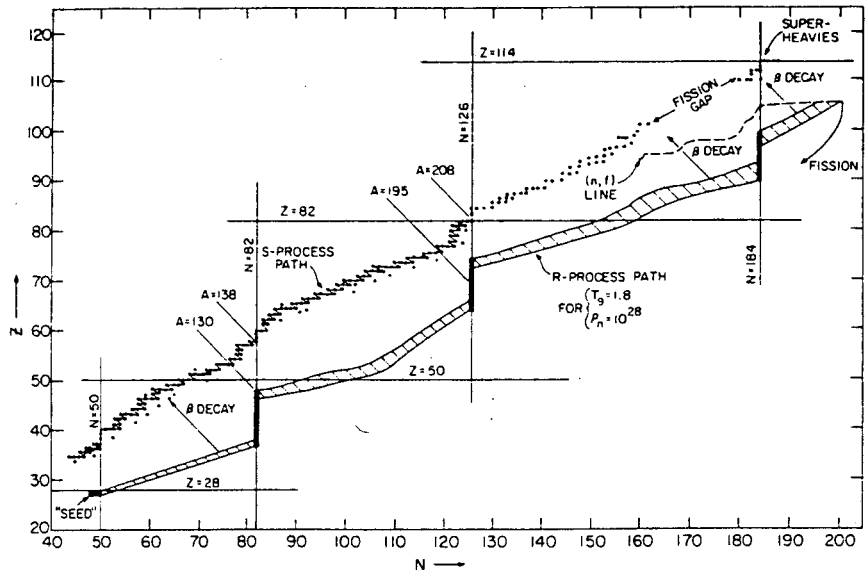

Figure 11. Neutron-capture paths for the s- and r-process [Ref. 20].

The same neutron magic numbers that lead to the production of s-process abundance peaks play a similar role in the r-process. Nuclei with neutron numbers 50, 82, and 126 have smaller than average neutron-capture cross sections and longer than average beta-decay half lives. Thus these nuclei become major r-process accumulation points. However, because of the huge neutron flux, these neutron magic numbers are reached at about 10 units lower atomic number (and hence atomic weight) in the r-process than in the s-process. In addition, because its path lies so far to the neutron-rich side of stability, the r-process avoids the region of alpha decays around bismuth. Thus, the r-process accounts for the abundance peaks seen at A = 80, 130, and 195, and the existence of the actinides.

In order to generate the high neutron flux needed for the r-process, it is believed that a high temperature environment is required. Under such conditions, the rate of change of the abundance of a particular nuclide (Z,N) is

$$\begin{aligned} dn(Z,N)/dt = & -\lambda_{n\gamma}(Z,N)n(Z,N) + \lambda_{n\gamma}(Z,N-1)n(Z,N-1) \\ & -\lambda_{\gamma n}(Z,N)n(Z,N) + \lambda_{\gamma n}(Z,N+1)n(Z,N+1) \\ & -\lambda_{\beta}(Z,N)n(Z,N) + \lambda_{\beta}(Z-1,N+1)n(Z-1,N+1) \end{aligned} \qquad (32)$$

where $\lambda_{n\gamma}$ = neutron capture rate, $\lambda_{\gamma n}$ = photoneutron reaction rate, λ_{β} = beta decay rate, and n(Z,N) = abundance of (Z,N). Thus, in order to perform r-process nucleosynthesis calculations, one needs estimates of many properties for a large number of neutron-rich nuclei. Neutron-capture cross sections, nuclear binding energies, and beta decay half lives are all important quantities in such calculations. To make such estimates more reliable, it is therefore important to extend measurements of these parameters as far from stability as possible.

In addition to the problems in r-process calculations due to uncertainties in the nuclear input data, the astronomical environment in which the r-process occurs remains an open question. A number of possible sites have been suggested: the regions just outside the cores of supernovae, supernova shocks, hydrodynamic instabilities of rotating magnetized stellar cores, and neutron star - black hole collisions. Despite these problems, a number of different r-process nucleosynthesis calculations have been performed. It has been found for certain assumed initial conditions, at least the qualitative features of the observed r-process abundances can be reproduced. Many details remain to be worked out, but the general outline of the r-process seems reasonably well established.

5. CONCLUSIONS

We have seen that charged-particle induced nuclear reactions are the sources of stellar energies. The ashes of these reactions are the elements between carbon and iron. The elements above iron are produced via neutron-capture reactions. Approximately half of the nuclides between iron and bismuth are synthesized in the s or slow process. The other half of these nuclides and all of the actinides are produced in the r or rapid process. The elements from carbon to uranium comprise only about 2% of the matter in the universe. The other 98% is in the form of hydrogen and helium, which are believed to have been produced in the Big Bang. The final fates of stars depend upon their masses. Low mass stars end their lives as white dwarfs that are supported against gravitational collapse by electron degeneracy pressure. More massive stars undergo supernova explosions which may leave behind a remnant. If the mass of the remnant

is $\lesssim 3\ M_\odot$ it will become a neutron star that is supported by neutron degeneracy pressure. Still more massive stars will collapse to become black holes.

6. ACKNOWLEDGEMENTS

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