Thus the calculated abundance is in fair agreement with that given by Suess and Urey. The expected decrease in cross sections with increasing N, neglected here, would give the observed gradual rise in abundance with N up to Hg^{202}. The drop in abundance at Hg^{204} is to be expected since it is not produced in the s process.

The large abundance of lead which we have derived from both the s and the r process results in considerable difficulties from the geological standpoint. Thus it is worthwhile to consider ways in which this conflict might be avoided. The most convincing evidence that the s process has been fully operative is the fact that the observed relative abundance of Pb^{208}, for which the r-process contribution is only 5%, and Pb^{204}, produced only by the s process, agree with the predictions of the s-process theory. Thus, the large Pb^{208}/Pb^{204} abundance ratio is attributable partly to the small neutron-capture cross section expected for magic Pb^{208} and partly to some cycling at the end of the s process. On the other hand, Pb^{204} might have an anomalously large cross section or might not be produced at all if Tl^{204} had such a large cross section that it captured a neutron and formed Tl^{205} instead of decaying to Pb^{204}. Neither of these possibilities seem likely, but they must be borne in mind as possible ways out of the conflict with geological evidence. It may be remarked that the total predicted lead production by the r process alone is 1.15 as compared with Suess and Urey's value of 0.47.

Finally, we consider the abundances of thorium and uranium. The recent results of Turkevich, Hamaguchi, and Reed (Tu56), using the neutron activation method, indicate that there is only a small amount of uranium in the Beddgelert chondrite. Urey (Ur56) has analyzed these results, which give an abundance of 0.007 on the scale of Suess and Urey. His results also give an atomic abundance of 0.02 for thorium. Our predicted r-process abundances are 0.147 for uranium and 0.462 for thorium. Thus, our calculations would seem to indicate that thorium and uranium as well as lead and bismuth have been reduced in abundance by some fractionation process in the formation of the planets and meteorites. As previously stated, of these four elements only lead has had its abundance in the sun determined (Go57), and in this case the solar abundance is in agreement with our calculated abundance and is very much higher than the abundance given by Suess and Urey.

In conclusion, the expected yield of radiogenic lead from the decay of thorium and uranium during the period since the formation of the meteorites would result in ratios of $(\mathrm{Rd}\,Pb^{206})/Pb^{204}$, $(\mathrm{Rd}\,Pb^{207})/Pb^{204}$, and $(\mathrm{Rd}\,Pb^{208})/Pb^{204}$ equal to 0.75, 0.41, and 0.57, respectively (see last line of Table VIII,4; Pb^{204} is taken as 0.2). These values are obtained by assuming no uranium-lead-thorium fractionation at formation. The Rd Pb^{206} and Rd Pb^{207} ratios are clearly consistent with Patterson's determinations of the radiogenic leads in chondrites and the Nuevo Laredo meteorite since we have used 4.5×10^9 years as the age of the meteorites. On the other hand, the ratios to Pb^{204} are considerably smaller than those found for radiogenic leads, indicating that lead was removed preferentially relative to uranium and thorium in chondrites and especially in the Nuevo Laredo meteorite. However, the iron meteorites could well contain the small amount of radiogenic leads which would be expected if no fractionation of lead relative to uranium and thorium occurred when they were formed.

IX. p PROCESS

In Sec. V it was pointed out that the proton-rich isotopes of a large number of heavy elements cannot be built by either the s or the r process, and particular examples of this difficulty are discussed in Sec. II, in connection with our assignments to one or the other of the synthesizing processes. With the exceptions of In^{113} and Sn^{115}, all of these isotopes have even A. From the Appendix we see that in the cases of Mo^{94}, Cd^{108}, and Gd^{152}, in addition to synthesis by the p process, these isotopes can also be made in weak loops of the s-process chain. In addition, In^{113}, Sn^{114}, and Sn^{115} may also be built in a weak loop of the s-process chain which is begun by decay from an isomeric state of Cd^{113}. In Fig. IX,1 a plot of the abundances of the p-process isotopes is given. The black dots represent isotopes which are made only by the p process, while the open circles represent isotopes which might have a contribution from the s process as well. A smooth curve has been drawn through the points merely to show the trends and positions of the peaks. This curve shows the same general trend as that of the main abundance curve, and the fact that the open circles lie predominately near the curve defined by the other points suggests that in no case is the s process an important contributor to the abundance. All of the isotopes are rare in comparison with the other isotopes of the same element, and it appears that only about 1% of the material has been processed by reactions which give rise to these isotopes (see Table II,1). The reactions which must be involved in synthesizing these isotopes are (p,γ) and possibly (γ,n) reactions on material which has already been synthesized by the s and the r processes.

The astrophysical circumstances in which these reactions can take place must be such that material of density $\gtrsim10^2$ g/cm^3 and containing a normal or excessive abundance of hydrogen is heated to temperatures of $2\text{–}3\times10^9$ degrees. It has been suggested (Bu56) that these conditions are reached in the envelope of a supernova of Type II. Alternatively they might be reached in the outermost parts of the envelope of a supernova of Type I in which the r process has taken place in the inner envelope. These possible situations are explored further in Sec. XII.

We suppose that for a short period quasi-statistical

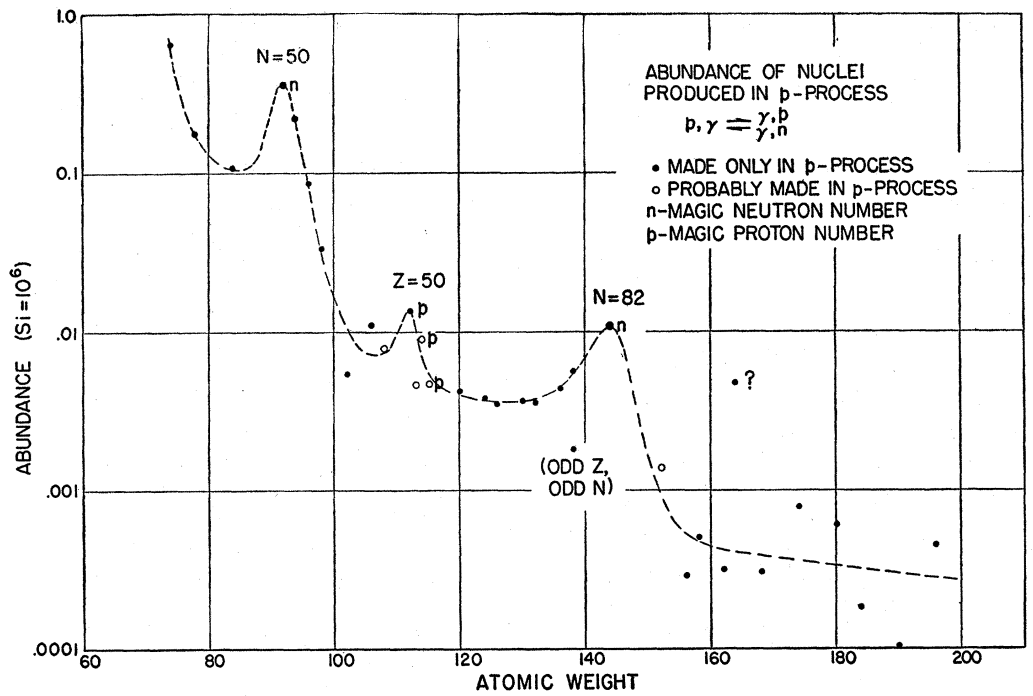

FIG. IX,1. Here we show a plot of the abundances of the isotopes made in the p process. The isotopes with magic N or magic Z are marked n and p, respectively. A curve has been drawn through the points to show the general trends. Note the peaks at $N=50$ and 82 and the lesser peak at $Z=50$.

equilibrium is reached between (p,γ), (γ,p), and (γ,n) reactions, i.e. $(p,\gamma)\rightleftarrows\begin{smallmatrix}(\gamma,p)\\(\gamma,n)\end{smallmatrix}$. Since the initiating reaction is (p,γ), the flux of free neutrons built up by any (γ,n) reactions will not become comparable with the proton flux, so that complete equilibrium cannot be set up between protons, neutrons, and gamma radiation. In Fig. IX,2 we give a schematic diagram which shows what the effect of (p,γ) and (γ,n) reactions on nuclei which originally lie on the stability curve in the Z, A plane will be. These reactions tend to drive nuclei off the curve of greatest stability in the direction of increasing A and Z in the case of (p,γ) reactions and decreasing A in the case of (γ,n) reactions. Qualitative arguments suggest that the values of ΔA and ΔZ, the displacements off the main stability curve, will be small. The reasons for this are as follows. In the cases of ruthenium, cadmium, xenon, barium, cerium, and dysprosium, the two lightest isotopes built by the p process have roughly equal abundances or at least ratios which never exceed ~ 3. In the case of tin, three isotopes, Sn^{112}, Sn^{114}, and Sn^{115}, have abundances in the ratio $\sim 3:2:1$. Now if it were supposed that (γ,n) reactions on the heavier isotopes were mainly responsible for the production of those proton-rich isotopes we might expect, because of the cross-section effect, that in the case of tin, for example, the ratio would be more nearly like $1:x:x^2$ where $x \gg 1$. If, on the other hand, we supposed that (p,γ) reactions were responsible for driving the nuclei a very long way from the main stability line so that contributions from many nuclei were responsible for the tin isotopes, we would also expect that the ratios would not be small. Thus our qualitative conclusion is that (p,γ) reactions are responsible but the nuclei are only displaced in general two or three units of A and Z from the stability line. The existence of the nuclei Mo^{92} and Sm^{144} which have closed shells of 50 and 82 neutrons respectively and which show up as peaks in Fig. IX,1 suggests that these have been made directly from progenitors with closed neutron shells which form peaks in the normal abundance curve; i.e. in the case of Mo^{92}, these would be Zr^{90}, Y^{89}, Sr^{88}, etc., while the progenitors of Sm^{144} would be Nd^{142}, Pr^{141}, Ce^{140}, etc. The case of Er^{164} remains an anomaly, since the ratio $Er^{164}/Er^{162} \simeq 15$ is very large and is out of line with the results for the other pairs of isotopes. There appear to be no peculiarities in the possible progenitors of Er^{164} which might explain this large abundance.

A simple calculation can be made to see whether these qualitative results deduced from the observed abundances are correct. The equation of statistical equilibrium is

$$\log\frac{n(A+1,Z+1)}{n(A,Z)}=\log n_p-34.07-\tfrac{3}{2}\log T_9+\frac{5.04}{T_9}Q_p, \quad (26)$$

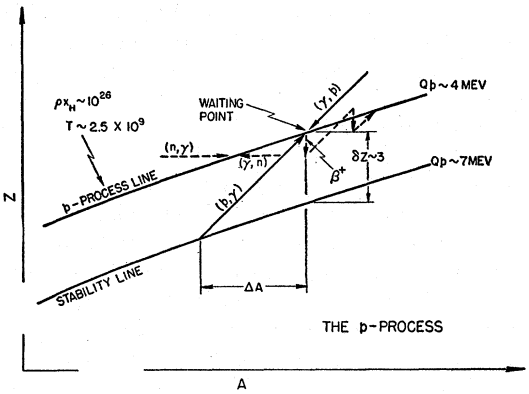

FIG. IX,2. The path of the p process in the Z, A plane. Material on the stability line (produced previously by the r or s process) is subjected to a hydrogen density of 10^{26} protons/cm^3 and a temperature of $\sim 2.5\times 10^9$ degrees. The (p,γ) reactions give an increase $\Delta A=\Delta Z=4$ or 5 until stopped by the (γ,p) reaction at nuclei with a proton binding energy, $Q_p\sim 4$ Mev. The (γ,n) reactions also produce a displacement onto the p-process line. Along this line positron emission must occur before further synthesis can take place. In general the lifetime of this emission is $\sim 10^3$ sec which is longer than the p-process conditions hold (100 sec). Thus a displacement $\delta Z\sim 3$ off the stability line (slope $\sim\frac{1}{3}$) at a given A occurs. This displacement by proton capture will occur even at large $A\sim 200$ in a time $\lesssim 1$ sec which is short compared to the time for the p process. Thus the displacement is terminated by the positron emission and not by decreasing Coulomb barrier penetrability so that it will be substantially independent of A. The p-process abundances as shown in Fig. I,1 and Fig. IX,1 should parallel the main abundance curve. The abundances will be 10^{-2} to 10^{-3} of those produced in the s and r processes.

where n_p is the number density of protons, Q_p is the proton binding energy, and T_9 is the temperature in units of 10^9 degrees. This equation is analogous to (14) in Sec. VII. Putting $n_p=10^{26}/\text{cm}^3$ and $T_9=2.5$, the binding energy of the last proton such that $n(A+1, Z+1)\approx n(A,Z)$ is obtained by putting the left-hand side of (26) equal to zero. We find $Q_p=4.3$ Mev. Now the proton binding energy is given by

$$Q_p(A,Z)=\alpha-\beta\left(1-\frac{4N^2}{Z^2}\right)-\tfrac{2}{3}\gamma A^{-\frac{1}{3}} -2\epsilon\frac{Z}{A^{\frac{1}{3}}}+\frac{\epsilon}{3}\frac{Z^2}{A^{\frac{1}{3}}}+g'(Z). \quad (27)$$

This equation is analogous to (23) of Sec. VII for the neutron binding energy; the symbols have the same meaning.

Thus

$$\left.\frac{\partial Q_p}{\partial A}\right|_N \approx -8\beta\frac{N^2}{A^3}\approx 0.6 \quad \text{for} \quad 100<A<200.$$

The binding energy of a proton in a nucleus on the main stability curve in this range of A is about 7 Mev. Thus $\Delta Q_p\simeq 2.7$ Mev and $\Delta A\simeq 4$ to 5. Thus ΔZ is also 4 to 5. However, we wish to know the deviation in Z of the new point from the stability line at $A+\Delta A$. We call this δZ and calculate it as follows. In Fig. IX,2 the slope of the line joining nuclei (A,Z) and $(A+1, Z+1)$ in the Z, A plane is 1. On the other hand, the change in Z corresponding to ΔA is just the change in the Z ordinate. Now the slope of the curve of maximum stability is $\sim\frac{1}{3}$. Thus δZ is given by

$$\delta Z=\tfrac{2}{3}\Delta A\simeq 3.$$

Whether the nuclei are driven off the main line to this maximum extent will depend on (i) whether sufficient protons are available, and (ii) whether the equilibrium conditions endure for sufficient time to allow the unstable nuclei to positron-decay so that the maximum number of protons can be added. The lifetime against positron emission is given by

$$\tau_{\beta^+}\simeq\frac{10^6}{W_{\beta^+}{}^5} \quad \text{(forbidden transition)},$$

where

$$W_{\beta^+}=B_A(\delta Z-0.5)+0.5.$$

Now $B_A\simeq 1.5$ Mev near $A=100$. The coefficient of B_A in this equation is $(\delta Z-0.5)$ instead of $(\delta Z-2.5)$ (which is used in Sec. VII) because we assume that the positron emission takes place by a forbidden transition to the ground state instead of by an allowed transition to an excited state as is the case for beta decay discussed in Sec. VII. In that case a mean value $\overline{W}_\beta$ was calculated.

Thus $W_\beta\simeq 4.2$, and $\tau_{\beta^+}\simeq 1000$ sec. Now the duration of a supernova outburst has been estimated to be 10–100 sec, and the time during which the p process takes place may be of the same order as or shorter than this explosion time. Consequently, it appears that the number of protons which can be added to the heavy nuclei is limited by the positron decay times and the synthesis to the limit of proton stability at this temperature will not be reached. Thus we conclude that the qualitative argument based on the observed abundances, that $\Delta A\sim 2$ or 3 and $\Delta Z\sim 1$ or 2, is borne out by this calculation.

The number of protons available to be captured by the heavy elements is determined by the number which are captured by the light and abundant elements. A typical example is C^{12}. Addition of two protons produces O^{14} whose half-life for positron emission is 72 sec. Thus it is probable that the maximum number of protons which can be added to C^{12} through the duration of the p process is only two. If we suppose that in the envelope in which the process occurs the H/C^{12} ratio is normal and equal to $\sim 10^4$, it is clear that even when we take into account all of the light elements which capture protons they can take only a small fraction of the total available, so that proton capture among the heavy nuclei is not limited by the number of protons available.

The mean reaction time for a (p,γ) reaction on a heavy nucleus A_0, Z_0 is given approximately by [see

Sec. III A, case (ii)]

$$\frac{1}{\tau}=3.1\times10^{9}\rho\frac{Z_0^{5/6}}{A_0^{1/6}T_9^{\frac{2}{3}}}$$

$$\times\exp\left[1.26\{AZ_0(A_0^{\frac{1}{3}}+1)\}^{\frac{1}{2}}-4.25\left(\frac{Z_0^2A}{T_9}\right)^{\frac{1}{3}}\right]\text{sec}^{-1},$$

where $A=A_0/(A_0+1)$. For example, for ${}_{80}Hg^{200}$, we have

$$\frac{1}{\tau}=3.0\times10^{23}\frac{\rho}{T_9^{\frac{2}{3}}}\exp(-78.8T_9^{-\frac{1}{3}})\text{ sec}^{-1}.$$

Thus for $T_9=2.5$ and $\rho=10^2$ g/cm^3, $1/\tau=10$ sec^{-1}, so that the capture rate is rapid and is not a limiting factor in the proton addition.

X. x PROCESS

We have given the name x process collectively to mechanisms which may synthesize deuterium, lithium, beryllium, and boron. Some discussion of the problems involved in the x process are discussed in this section.

A. Observational Evidence for the Presence of Deuterium, Lithium, Beryllium, and Boron in our Galaxy

From the appendix the number ratios of deuterium, Li^6, Li^7, Be^9, B^{10}, and B^{11} relative to hydrogen are 1.4×10^{-4}, 1.8×10^{-10}, 2.3×10^{-9}, 5×10^{-10}, 1.1×10^{-10}, and 4.9×10^{-10}, respectively. Thus deuterium is rare as compared with its neighbors hydrogen and helium in the atomic abundance table, but as far as the remainder of the elements are concerned is very abundant and comparable with iron. Lithium, beryllium, and boron are all extremely rare as compared with their neighbors helium, carbon, nitrogen, and oxygen, and are only about 100 times as abundant as the majority of the heavy elements (see Fig. I,1). It should be emphasized for these elements particularly that all of these values have been obtained from terrestrial and meteoritic data and thus they may not be at all representative of the cosmic abundances of these elements.

A number of attempts have been made to detect deuterium in the sun. The latest, by Kinman (Ki56), shows that the abundance ratio of deuterium to hydrogen in the atmosphere is less than 4×10^{-5}. Attempts have also been made to detect the radio spectral line at 327 Mc/sec due to neutral deuterium in interstellar gas, and the most recent results by Stanley and Price (St56) and Adgie and Hey (Ad57) lead to the conclusion that the deuterium to hydrogen ratio does not exceed 5×10^{-4}.

The abundances of lithium and beryllium in the solar atmosphere have been investigated by Greenstein and Richardson (Gr51) and Greenstein and Tandberg-Hanssen (Gr54c). They found that while the isotope ratio Li^6/Li^7 could well be normal, the lithium to hydrogen ratio lies in the range $1.6\times10^{-11}-7\times10^{-12}$ or about 100 times smaller than the meteoritic value. The beryllium to hydrogen ratio$=10^{-10}$, a value which is reasonably well in agreement with that of Suess and Urey. An upper limit to the beryllium to hydrogen ratio of 10^{-10} has been obtained for a magnetic star by Fowler, Burbidge, and Burbidge (Fo55a). Estimates of the upper limits to the interstellar abundances of lithium and beryllium have been made by Spitzer (Sp49, Sp55). He has estimated that the upper limit to the lithium to sodium ratio is about 0.1, so that if sodium has normal abundance the lithium to hydrogen ratio is less than 10^{-7}. The beryllium to hydrogen ratio is found to be $\leq10^{-11}$. This is an order of magnitude lower than that found in meteorites, and Spitzer has concluded that some beryllium may be locked up in interstellar grains.

Lithium, probably in variable amounts, has been detected through the presence of the Li I resonance doublet at λ 6707.8 in a wide variety of late-type stars. Certain carbon stars show this feature very strongly in their spectra (Mc40, Mc41, Mc44, Sa44, Fu56), while it is also present in S-stars (Ke56, Te56), and in M-type dwarfs and giants (Gr57). No relative abundances are yet available.

Boron is spectroscopically unobservable.

In the primary cosmic radiation the abundances of lithium, beryllium, and boron are comparable with those of carbon, nitrogen, and oxygen (Go54a, Ka54, No55, No57).

B. Nuclear Reactions Which Destroy Deuterium, Lithium, Beryllium, and Boron

First we outline the reactions which destroy these light elements in a hydrogen-burning zone. At temperatures commonly found in the interiors of main-sequence stars the reactions which destroy deuterium are $D^2(d,p)H^3(\beta^-)He^3$, $D^2(p,\gamma)He^3$, and $D^2(d,n)He^3$. In the same way the reactions which destroy lithium are $Li^6(p,\alpha)He^3$ and $Li^7(p,\alpha)He^4$. Beryllium is destroyed by $Be^9(p,d)Be^8\rightarrow2He^4$ and $D^2(p,\gamma)He^3$, or $Be^9(p,\alpha)Li^6(p,\alpha)He^3$. Boron is destroyed by $B^{10}(p,\alpha)Be^7(EC)$-$Li^7(p,\alpha)He^4$ and $B^{11}(p,\alpha)Be^8\rightarrow2He^4$. Thus the net result is always to convert these elements into helium through proton bombardment, and the rates of the reactions are such that in all conditions before a star evolves off the main sequence all of the deuterium, lithium, beryllium, and boron in the volume which contains the vast majority of the mass will be destroyed (Sa55).

C. Synthesis of Deuterium, Lithium, Beryllium, and Boron

The foregoing considerations make it appear probable that these elements have been synthesized in a low-temperature, low-density environment in the universe, or conceivably in a region in which hydrogen was

absent. The alternative is a situation in which the synthesis was followed extremely rapidly by an expansion of the material resulting in cooling and decreased density so that the reactions leading to destruction were avoided, or else by a rapid transfer of the synthesized material to a low-temperature, low-density environment. Recent work of Heller (He57) shows very clearly the conditions of density and temperature under which a deuterium to hydrogen ratio $\simeq 10^{-4}$ can be preserved. Now the regions in our Galaxy in which such conditions are present are (i) in stellar atmospheres and (ii) in gaseous nebulae.

(i) This possibility was explored (Fo55a, Bu57) in connection with nuclear reactions which may take place in the atmospheres of magnetic stars (see Sec. XI). Besides having large magnetic fields and thus large sources of magnetic energy which may be available for particle acceleration and hence nuclear activity, the magnetic stars have one other feature which makes the results of elements synthesis in their atmospheres observable. It seems probable that their atmospheres do not mix appreciably with the lower layers. Thus the light elements synthesized there will not be mixed into the interiors and destroyed. However, the abundance of deuterium produced following the release of neutrons by (p,n) reactions on the light elements in such an atmosphere can never exceed a deuterium to hydrogen ratio $\sim 10^{-2}$. The use of different nuclear cross sections than those employed in Fo55a might conceivably raise this upper limit to $\sim 10^{-1}$. Because of the rarity of the magnetic stars and because of the small mass of their atmospheres, ejection of such atmospheres into the interstellar gas will lead to an interstellar deuterium to hydrogen ratio $\leq 10^{-10}$. If a more widespread class of stars such as the red dwarfs are proposed as stars in which deuterium can be made in the same way, a maximum interstellar ratio of 10^{-6} might be achieved. Thus since the lower limit to the interstellar abundance ratio so far determined is 5×10^{-4}, this explanation might be satisfactory. However, it leaves the terrestrial abundance to be explained by circumstances peculiar to the solar system, e.g., through electromagnetic activity in an early stage of formation of the solar nebula, a situation which is rather unsatisfactory.

Production of lithium, beryllium, and boron in a stellar atmosphere can take place through spallation reactions on abundant elements such as carbon, nitrogen, oxygen, and iron. Thus, if we believe that stellar atmospheres are the places of origin of these elements, it is also probable that they are a major source of the primary cosmic radiation, a conclusion which is consistent with observed abundances of primary nuclei mentioned earlier. Since energies $\gtrsim 100$ Mev/nucleon are demanded for spallation reactions it has been suggested that these reactions take place in regions fairly high in the stellar atmospheres. If it is supposed that these reactions go on only in magnetic stars, and that all the material synthesized is ejected, then the interstellar ratios of lithium, beryllium, and boron to hydrogen which can be obtained are all about 10^{-18}. However, if the red dwarfs are proposed as stars in whose atmospheres spallation reactions can go on at the same rates, ratios $\sim 10^{-10}$ can reasonably be predicted for the interstellar gas. Lithium has been found to be quite strong in the spectrum of T Tauri (He57a).

Thus we conclude that if the x process takes place in stellar atmospheres, a further process is demanded to synthesize the deuterium in the solar system, though the interstellar abundances of lithium, beryllium, and boron might be produced.

(ii) We shall now consider in a qualitative fashion the possibility that the x process takes place at some point in the evolution of gaseous nebulae. The current sources of energy in the bright gaseous nebulae are in general the highly luminous stars which are embedded in them. However, the discovery in recent years that some gaseous nebulae (not necessarily optically bright) are strong sources of nonthermal radio emission has shown that another source of energy must be present. There is strong support for the theory that this radiation is synchrotron emission by high-energy electrons and positrons moving in magnetic fields in these sources. The reservoirs of energy which must be present in such sources in the form of *high-energy particles and magnetic flux* are as large as the total amounts of energy which are released in supernova outbursts. Thus the energetic conditions that are demanded to produce a flux of free neutrons, which can then be captured to form deuterium and can also spallate nuclei to form lithium, beryllium, and boron, are also present. Unfortunately, the mean densities in such radio sources are so low ($\sim 10^{-21}$ g/cc) that at the present time the amount of activity leading to synthesis must be negligible. On the other hand, the situation in the early life-history of such nebulae may have been more conducive to such synthesis.

Recently the light from another star of the T Tauri class, NX Monocerotis, has been found to be largely polarized, thus indicating the presence of a considerable amount of synchrotron radiation (Hu57c). High-energy particles and a magnetic field must therefore be present. Stars in the T Tauri class are thought to be young stars, recently formed and not yet stabilized on the main sequence, and are embedded in dense interstellar matter. The observation of lithium in T Tauri itself and synchrotron radiation in NX Monocerotis support the idea that the x process (by spallation) is occurring here.

Of the strong radio emitters in our own galaxy, the Crab is known to be a supernova remnant, and it may be that some of the other sources are supernova remnants as well. In the context of the theory described in this paper, supernovae are the sources both of the r process and the p process, while a supernova outburst may well follow the e-process production of elements. Now if, in the early stages of a supernova outburst, after these other processes have taken place and the

envelope is beginning to expand so that the densities fall below those in which deuterium, for example. will be destroyed, the envelope can be subjected to a large flux of neutrons, it is possible that the x process could take place. However, as pointed out by Heller, if the source of neutrons which is demanded to synthesize deuterium is $C^{13}(\alpha,n)O^{16}$ or $Ne^{21}(\alpha,n)Mg^{24}$ (in Sec. III it was pointed out that there are strong reasons for rejecting $C^{13}(\alpha,n)O^{16}$ and using $Ne^{21}(\alpha,n)Mg^{24}$ as the source of neutrons here), then a high temperature $\sim 10^9$ degrees is demanded to produce sufficient neutrons, but a low density $\sim 10^{-4}-10^{-5}$ g/cm^3 is demanded to preserve the deuterium; this appears to be an improbable situation. Thus a more plausible mode of synthesis would obtain if a large flux of neutrons from a high-energy, high-temperature region could be injected into an envelope of cool hydrogen at an early stage of the expansion. The following situation may be envisaged. As discussed in Sec. XII, at temperatures in excess of 5×10^9 degrees and densities $\geq 10^8$ g/cm^3 the stellar core will reach a configuration in which there can easily be a transition from a core made of iron to a core consisting primarily of helium with a fraction of free neutrons. If some portion of this core can be ejected into a surrounding, relatively cool, shell, the neutrons will be moderated and captured to form deuterium. The condition that they are captured before decaying is that $\bar{\rho}\geq 10^{-10}$ g/cm^3.

The very advanced stage of evolution in which a star might possess a neutron core might also provide a suitable source of neutrons for deuterium production, though this question will not be explored further here.

In Sec. XII it is shown that one percent of a supernova shell may be converted to the heavy elements in the r process. If the total mass of material is taken to be $\sim 1M_\odot$, then dilution by a factor of the order of 10^4 is demanded to produce the correct abundances, relative to hydrogen, in the solar system, which is $\sim 10^{-6}$ from Table II,1. Thus, the supernova shell must have mixed with a mass of about 10^4 times its mass in the interstellar gas. The deuterium to hydrogen ratio in the solar system is $\sim 10^{-4}$ so it appears that if sufficient neutrons were available to convert the whole of the hydrogen in a supernova shell to deuterium, then this amount of dilution could explain the observed solar system abundance of this isotope. To produce the lithium, beryllium, and boron, it must be assumed that about 1% of the carbon, nitrogen, and oxygen in the shell was spallated either by high-energy neutrons or alpha particles ejected by the core or by protons into which the neutrons decayed. It is also possible that these latter elements may have been produced by a high-energy tail of the protons which are demanded in the p process. It does not appear probable, however, that a single supernova could account both for the deuterium and the r process elements in the solar system.

A supernova origin for the x-process elements would also suggest that supernovae are the original sources of cosmic radiation. Thus the primary cosmic radiation might be expected to have greater than normal abundances of iron and other abundant elements synthesized in supernovae. There is some evidence in support of this (No57).

D. Preservation of Lithium in Stars

Finally, we discuss an observational result which may suggest yet another region where lithium may be synthesized, and where it is certainly preserved. We summarized in Sec. X A the evidence that lithium is present in many kinds of late-type stars. Though no relative abundances of this element in different classes of late-type stars have yet been determined, it seems highly probable that the lithium abundance is variable, and also in some stars it is probably far more abundant than it is in the interstellar gas. This suggests strongly that there is some mechanism by which it can be preserved and even synthesized in stellar interiors. For example, in one S-type star, R Andromedae (Me56), both technetium and lithium are observed. Hence in this case there must be mixing between the helium-burning core where the technetium is synthesized and the surface. We have emphasized in Sec. X B that lithium cannot survive in a hydrogen-burning zone. Cameron (Ca55) attempted to show that Li^7 can be built in a hydrogen-burning zone through $He^3(\alpha,\gamma)Be^7(K)Li^7$, the Li^7 reaching the surface of the star because it is preserved in the form of Be^7, whose half-life against K capture is lengthened because of the comparative rarity of s-state electrons at the densities of production. Unfortunately his calculation of the rate of the reaction $He^3(\alpha,\gamma)Be^7$ is based on an estimate given by Salpeter (Sa52a) which was incorrectly printed, and it seems also that the increase of the half-life against K capture has been considerably overestimated.

Lithium does not react as rapidly with helium as it does with hydrogen, and it appears possible that it can exist in a helium-burning zone long enough for it to be able to reach the surface from the deep interior. Thus we conclude that the presence of lithium in the atmospheres of late-type stars suggests that these stars can have no hydrogen-burning zones. This implies that the helium cores must be large enough, and the stars' evolution sufficiently advanced, so that the hydrogen envelopes extend only to depths where the temperatures are less than $\sim 10^6$ degrees. Reactions which may synthesize lithium in a helium-burning core will be discussed elsewhere.

XI. VARIATIONS IN CHEMICAL COMPOSITION AMONG STARS, AND THEIR BEARING ON THE VARIOUS SYNTHESIZING PROCESSES

Different processes of element synthesis take place at different epochs in the life-history of a star. Thus

the problem of element synthesis is closely allied to the problem of stellar evolution. In the last few years work from both the observational and theoretical sides has led to a considerable advance in our knowledge of stellar evolution. Theoretical work by Hoyle and Schwarzschild (Ho55) has been repeatedly mentioned since it affords estimates of temperatures in the helium core and in the hydrogen-burning shell of a star in the red-giant stage. However, the calculated model is for a star of mass 1.2 solar masses and a low metal content (~1/20 of the abundances given by Suess and Urey), and is thus intended to apply to Population II stars.

From the observational side, the attack has been through photometric observations of clusters of stars, from which their luminosities and surface temperatures have been plotted in color-magnitude or Hertzsprung-Russell (HR) diagrams, by many workers. We refer here only to two studies, by Johnson (Jo54) and by Sandage (Sa57a). A cluster may be assumed to consist of stars of approximately the same age and initial composition, and hence its HR diagram represents a "snapshot" of the stage to which evolution has carried its more massive members, once they have started to evolve fairly rapidly off the main sequence after their helium cores have grown to contain 10 to 30% of the stellar mass. Clusters of different ages will have main sequences extending upwards to stars of different luminosity, and the point at which its main sequence ends can be used to date a cluster; the rate of use of nuclear fuel, given by the luminosity, depends on the mass raised to a fairly high power (3 or 4).

Figure XI,1 is due to Sandage (Sa57a), and is a composite HR diagram of a number of galactic (Population I) clusters together with one globular cluster, M3 (Population II). The right-hand ordinate gives the ages corresponding to main sequences extending upwards to given luminosities. This diagram gives an idea of the way in which stars of different masses (which determine their position on the main sequence) evolve into the red-giant region. The most massive stars evolve into red supergiants (e.g., the cluster h and χ Persei); the difference between the red giants belonging to the Population I cluster M67 and the Population II cluster M3 should also be noted. Since stars evolve quite rapidly, compared with the time they spend on the main sequence, the observed HR diagram may be taken as nearly representing the actual evolutionary tracks in the luminosity-surface-temperature plane.

A feature of Population II stellar systems is that their HR diagrams contain a "horizontal branch" [see, for example, the work by Arp, Baum, and Sandage (Ar53)] which probably represents their evolutionary path subsequent to the red-giant stage. This feature is not included in the diagram of M3 in Fig. XI,1, although it is well represented by that cluster, since this diagram is intended to show just the red-giant branches of clusters. The old Population I cluster M67 has a sparse distribution of stars which may lie on the Population I

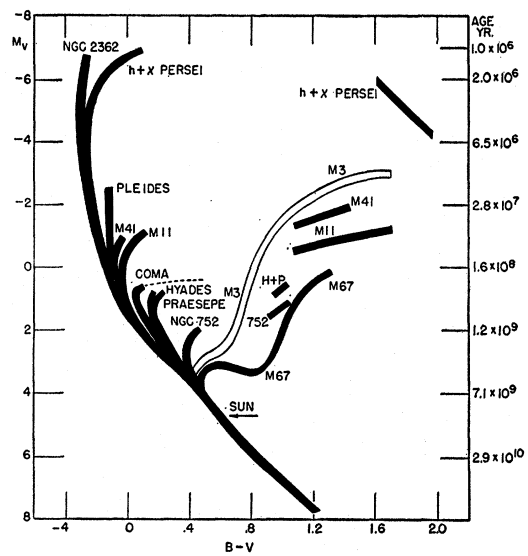

FIG. XI,1. Composite Hertzsprung-Russell diagram of a number of galactic (Population I) star clusters, together with one globular cluster, M3 (Population II), by Sandage (Sa57a). The abscissa measures the color on the B-V system, and defines surface temperature (increasing from right to left). The left-hand ordinate gives the absolute visual magnitude, M_v, of the stars (thus luminosity increases upwards). The heavy black bands (Population I) and unfilled band (Population II) represent the regions in the temperature-luminosity plane which are occupied by stars. The names of each cluster are shown alongside the appropriate band. The right-hand ordinate gives the ages of the clusters, corresponding to main sequences extending upwards to given luminosities. Note that the clusters all have a common main sequence below about $M_v=+3.5$. Note also that the red giants have different luminosities, according to the luminosities they had while on the main sequence (which are defined by their masses). *Diagram reproduced by courtesy of the Astrophysical Journal.*

analogy of the horizontal branch; the other Population I clusters in Fig. XI,1 do not show such a feature. Presumably the evolutionary history of a more massive Population I star subsequent to its existence as a a red giant is more rapid.

Whenever reference is made in different parts of this paper to particular epochs in a star's evolutionary life, we are referring to a schematic evolutionary diagram for the star which has the same general characteristics as the HR diagrams in Fig. XI,1. With this background in mind, we turn now to astrophysical observations which provide many indications of element synthesis in stars. This is either taking place at the present time or else it has occurred over a time-scale spanned by the ages of nearby Population II stars. These indications may be divided into the following groups.

A. Hydrogen Burning and Helium Burning

It appears theoretically probable that some stars, perhaps delineated by their initial mass and amount of mass loss during their existence on the red-giant branch of their evolutionary path, may still exist as stable configurations for a time after they have ex-

hausted almost all of their hydrogen. Such stars would be expected to be very rich in helium and perhaps the products of helium burning. Alternatively, stars may develop inner cores in which helium burning takes place, while still possessing hydrogen-burning shells and envelopes with a normal hydrogen content. If mixing (large-scale convection) then sets in, in such a star, core material which has been modified by hydrogen burning may appear on the stellar surface. While a star is in the red-giant stage, the temperature of the hydrogen-burning zone may be as high as 30 or 40×10^6 degrees, at least in Population II stars with masses of 1.2 $M_\odot$. At 30×10^6 degrees there may be considerable destruction of oxygen through the reactions $O^{16}(p,\gamma)F^{17}(\beta^+\nu_+)O^{17}(p,\alpha)N^{14}$ (cf. Table III,2 and Fig. III,2), if the evolutionary time-scale is as long as 10^5 years at this stage. If the onset of energy generation by helium burning in a partially degenerate core leads to some instability and consequent mixing (Ho55), then the star may settle down into a new structure in which hydrogen burning occurs at a lower temperature, in material which may or may not have been appreciably enriched in C^{12}.

As was pointed out in Sec. III F(1), any stars in which the products of helium burning mix through a hydrogen-burning zone to the surface should appear rich in nitrogen and neon, as well as in helium, and considerably depleted in hydrogen. The equilibrium abundances of C^{13}, N^{14}, and N^{15} relative to C^{12} will be approached if sufficient time is spent in the hydrogen zone for the CN reactions to have gone through several cycles. On the basis of the most recent cross-section values N^{14} will be by far the most abundant of the carbon and nitrogen isotopes if equilibrium is attained even though C^{12} is the nucleus produced in the helium burning. O^{16} will also be converted to N^{14} (without cycling) if the temperature is great enough. Ne^{20} will mainly be converted to Ne^{21} and Ne^{22} and a very small amount of Na so that the element neon survives hydrogen-burning.

Greenstein (Gr54a), using the older rates of the CN cycle, discussed the connection between the operation of the CN cycle at various temperatures, the amount of mixing, and the abundance of nitrogen observed on the surface. We wish to elaborate on this argument at this point. If equilibrium is attained the ratio N/C will be given by 0.82 N^{14}/C^{12} since N^{15} is very rare and $C^{12}/C^{13}=4.6$ by number independently of temperature. The rough temperature dependence given for N^{14}/C^{12} in Sec. III F(1) is not sufficiently accurate for our present purposes and in Table XI,1 we give various equilibrium ratios as a function of temperature. In the last row of the table we list the logarithm of the time t in years for the N/C ratio to approach the equilibrium value at a given temperature from a not too greatly different value at another temperature. This time is of the order of magnitude of the sum of the mean lifetimes of C^{12} and C^{13} which is 22% greater than that of C^{12} alone and is thus easily calculable from the entries in Table III,2. In Table XI,1 we have taken $\rho x_H=10$ g/cc since this value is representative of the density and hydrogen concentrations in hydrogen-burning zones in giant stars. These characteristic times required to reach the equilibrium ratio for N/C are relevant to our problem. Only if the CN isotopes remain at the temperatures listed in the first row of Table XI,1 for times comparable to those listed in the last row will the N/C ratios given in the intervening rows be reached. The last and thus lowest temperature at which the CN isotopes spend sufficient time in mixing to the surface will determine the observed surface ratios. This assumes that the mixed-in CN outweighs that in the original envelope material. An inspection of the table indicates that this last, lowest temperature is probably about 15×10^6 degrees since the time required for N/C to reach its equilibrium value is $\sim10^7$ years at this temperature. We estimate that 10^7 years is the maximum time available for transport of the original C^{12} from the core through a significant portion of the envelope.

Figure XI,2 gives a schematic representation of the way in which the N/C ratio varies in the hydrogen-burning regions of a star of approximately four solar masses. During the initial condensation of the star, this ratio might be about 1 or 2, as in the solar system material and in young stars like 10 Lacertae and τ Scorpii (see Table XI,2). As hydrogen burning becomes established at central temperatures $20\text{–}30\times10^6$ degrees, the ratio rapidly reaches an equilibrium value in the range 100–50. When the star leaves the main sequence, gravitational contraction of the core occurs, the temperature of the hydrogen-burning shell rises and N/C decreases. At the onset of helium-burning, C^{12} is produced in the core and eventually some of this new carbon will be mixed into the hydrogen-burning regions. There will be a relatively short period during which the N/C ratio decreases to a minimum, followed by a rapid rise as N/C reaches its equilibrium value at the temperature at which the hydrogen burns in the star, which by now is in a highly evolved stage, per-

TABLE XI,1. Equilibrium nitrogen-carbon ratios as a function of temperature ($\rho x_H=10$).

T(10^6 degrees)	10	12	15	20	30	40	50	70	100
N^{14}/C^{12} by number	470	250	200	110	58	38	28	17	14
N/C by number	380	210	160	90	47	31	23	14	12
N/C by mass	440	240	180	100	54	36	26	16	13
$\text{Log}_{10}t$ (years)	10.6	9.0	7.2	5.1	2.4	0.8	−0.4	−2.1	−3.6

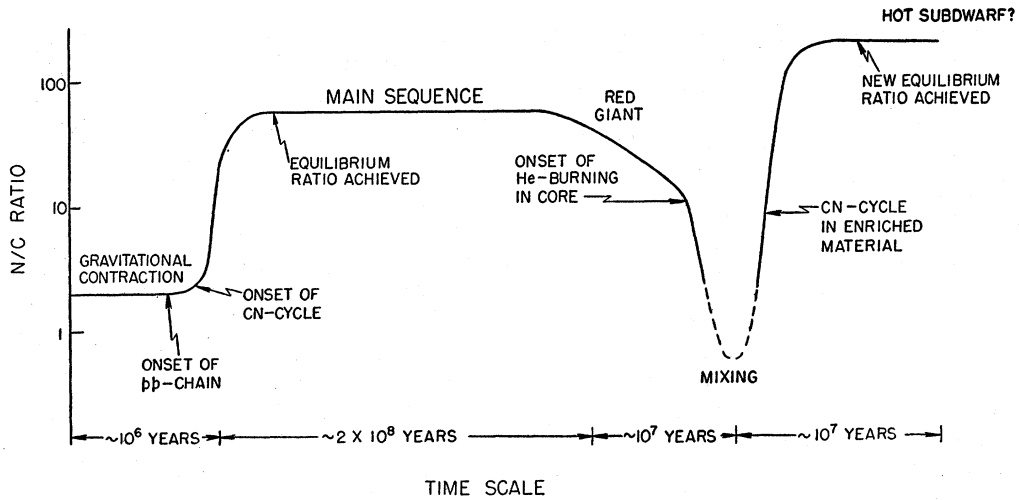

FIG. XI,2. Schematic representation of the variation of the N/C ratio in regions where the hydrogen burning occurs in a star of approximately four solar masses. An initial ratio of ~2, as in the solar system, is assumed. The time scales for different phases of the star's evolution are schematically marked as abscissas. With the onset of the CN cycle, an equilibrium ratio in the range ~50–100 is rapidly reached; when the star leaves the main sequence the temperature increases and the ratio drops. The onset of helium burning in the core reduces the ratio to an unknown lower limit defined by the onset of mixing. Passage of the core C^{12} through an outer hydrogen-burning region increases the ratio again to a new equilibrium value.

haps that of a hot subdwarf. The hydrogen-burning zone may be far out from the star's center and thus at a fairly low temperature, say 15×10^6 degrees. In this case we would estimate N/C to be ~160 with an upper limit from uncertainties in cross sections of ~250.

Several groups of stars exist whose spectra give evidence that hydrogen burning and helium burning have been occurring under the various possible conditions that we have just considered. These are as follows:

(i) Some stars, classified by other criteria as being of spectral types *O* or *B*, have no hydrogen lines; the stars HD 124448 (Po47), HD 160641 (Bi52), and HD 168476 (Th54) are examples. Helium and carbon lines are strong in all three; the oxygen lines usually prominent at this temperature are completely lacking in HD 168476. A preliminary determination of abundances in HD 160641 (Al54) has shown that carbon, nitrogen, and neon are all more abundant, relative to oxygen, than in normal stars, while helium has completely replaced hydrogen. These three stars have high velocities and may have originated in another part of the Galaxy than the solar neighborhood; they are probably Population II objects. The comparative rarity of such stars among surveys of *B*-type stars seems to indicate that they do not spend long in this evolutionary stage, or else the conditions under which stars evolve in this way are actually rare.

(ii) Münch and Greenstein have studied seven hot subdwarf stars with peculiar abundances of the light elements, and especially with an apparent excess of nitrogen. An analysis by Münch (Mu57a) of one such star, HZ 44, which is a high-temperature subdwarf, has shown that helium is more abundant than hydrogen. Nitrogen is about 200–300 times as abundant as carbon. Two determinations of the relative abundances of the elements, by mass, are given in Table XI,2, corresponding to the range of temperature and pressure given by the analysis. Aller's results for HD 160641 [see (i) above], are also given, together with the "normal" abundances by Suess and Urey, the abundances in the main-sequence (unevolved) young stars τ Scorpii and 10 Lacertae obtained by Traving (Tr55, Tr57), and the abundances in τ Scorpii obtained by Aller, Elste, and Jugaku (Al57c) (we have arbitrarily assumed here that the H/He ratio is the same as that determined by Traving). Since numbers of atoms are not conserved during element synthesis ($4H^1\rightarrow He^4$, $3He^4\rightarrow C^{12}$ etc.), only percentage abundances by mass are meaningful in considering the amount of increase or decrease in a given element, as compared with its normal abundance. In HZ 44 there appears to have been a slight increase in the sum of the elements heavier than helium, most of which is due to an increase in nitrogen by a factor between 10 and 25. The earlier discussion indi-

TABLE XI,2. A comparison between the chemical compositions of evolved stars and young stars.

	Percentage mass				
		Main sequence young stars			
Element	Suess and Urey (Su56)	Traving	Aller	HZ 44 (Range)	HD 160641
H	75.5	58.6	[58.7]	9.0 – 9.1	0
He	23.3	39.8	[39.8]	88.7 –83.8	95.4
C	0.080	0.17	0.03	0.0047– 0.014	0.31
N	0.17	0.19	0.15	1.66 – 4.85	0.47
O	0.65	0.55	0.40	0.12 – 0.34	0.75
Ne	0.32	0.62	0.84	0.47 – 1.38	3.00
Si	0.053	0.10	0.07	0.17 – 0.48	0.065

cates that this increase is probably due to modification of C^{12}, part of which was present in the star when it condensed, and part of which was produced by helium burning. The following schematic picture may explain the observations. The star has had a helium-burning core in which at least some C^{12} was built, and probably some O^{16} and Ne^{20} (depending on the time-scale and temperature, according to Fig. III,3). Mixing then occurred, and the core material passed out through a hot hydrogen-burning region (30–40×10^6 degrees) in which a good proportion of the oxygen could have been destroyed (depending on the time the material spent at this temperature). The neon isotopes cycled and were not destroyed even if the temperature was high enough for neon to interact. Further modification of the star's structure occurred, and the most recent hydrogen-burning region through which the internal material passed, on being mixed to the surface, was a cool one ($\sim15\times10^6$ degrees), in which the ratio N/C in the enriched material became ~160 by number on the basis of our previous equilibrium estimates and ~300 by observation. Since each of these values has errors of the order of $\pm50\%$ it would seem that they are in satisfactory agreement.

Münch has suggested that HZ 44 (and some other similar stars which he is studying at present) may be of Population I; they have low velocities and differ in other respects from "horizontal branch" stars of Population II, and may be analogous objects in Population I. The absolute magnitude of HZ 44 is in the range $+3$ to $+5$; the surface gravity seems high, and its mass may be considerably larger than 1 $M_\odot$.

(iii) The "classical" Wolf-Rayet stars, of both the carbon and the nitrogen groups, belong to Population I. A discussion, which includes references to earlier work, is given by Aller (Al43). There is still disagreement on the extent to which the characteristic broad emission features in the spectra are due to ejection of material, but there seems no doubt that these stars have reached a late evolutionary stage (Sa53). One example is a member of the binary system V 444 Cygni, in which the other component is an early-type massive star. The mass of the WR component is apparently less than the main-sequence component, but is still large (~10 $M_\odot$); if the star has reached a later evolutionary stage then its mass may originally have been larger. The lifetime on the main sequence may have been only a few million years. The WR stars are apparently deficient in hydrogen, but upper limits to its abundance have not been set. Helium is apparently the main constituent. In the WC stars the ratios by number of helium: carbon: oxygen are 17:3:1: (Al57a), and nitrogen is not seen at all. In the WN stars the ratios by number of helium: carbon: nitrogen: are about 20:1/20:1 (Al43, Al57a). All these estimates are rough, because of difficulties in abundance determinations, arising through large departures from thermodynamic equilibrium, stratification, etc. However, it seems probable that in WN stars the surface material has been through the CN cycle (Ga43); according to the dependence of N^{14}/C^{12} as a function of T, the temperature would be quite high, $\sim50\times10^6$ degrees. In WC stars, on the other hand, the products of helium burning have probably reached the surface without being modified by hydrogen burning (Sa53, Gr54a).

(iv) Planetary nebulae are Population II objects (Mi50). The composition of the nebulae is quite similar to that of young stars (Al57d), but the central stars, which can be of WR, *Of*, or absorption-line *O* type, are often apparently deficient in hydrogen. Many authors (e.g., Sw52) have pointed out that, among the central stars of WR type, there is not a clear-cut distinction between a carbon and a nitrogen sequence, as there is in Population I WR stars. Apparently the abundance ratio of carbon/nitrogen can take a whole range of values: the balance between the onset of mixing and the extent and temperature of a hydrogen-burning zone would seem not to be so critical as in the case of the Population I WR stars. The masses of the central stars of planetary nebulae are about 1 $M_\odot$ (Al56), and their absolute magnitudes are about 0 or $+1$. Bidelman (Bi57) and Herbig (St57) have suggested that the hydrogen-poor carbon stars of the R Coronae Borealis class [see (viii) below] may be the ancestors of some planetary nebulae.

(v) The *A*-type stars υ Sagittarii (Gr40, Gr47) and HD 30353 (Bi50) have anomalously weak hydrogen lines and appear to be rather similar to, although slightly cooler than, HZ 44 [see (ii)]. While helium, nitrogen, and neon are all strong in υ Sagittarii, carbon is weak. In Plate 1 (see pages 611–614 for plates) spectra are shown of the region near the Balmer limit in this star and in the *A*-type supergiant η Leonis (which it resembles in temperature and pressure), obtained by Greenstein at the McDonald Observatory. The comparison between the two stars is striking, the remarkable absence of the Balmer discontinuity in υ Sagittarii, and the strength of helium lines while at the same time ionized metals are strong, are impossible to explain other than by an abnormal composition. The great strength of lines other than hydrogen has been explained by the reduced opacity (mainly due to hydrogen in a stellar atmosphere of this temperature), but possibly the temperature is somewhat lower than has been estimated. Both υ Sagittarii and HD 30353 are spectroscopic binaries; this might be significant in view of work by Struve and Huang on mass-loss from close binaries (St57a), since one possibility for their evolutionary history is that the stars might have lost most of their hydrogen-containing atmospheres in this way.

(vi) Some white dwarfs, i.e., stars near the end of their life, show strong lines of helium and no hydrogen. These stars must have degenerate cores and the only nuclear fuel remaining must be in the surface layers. In Plate 1 we show Greenstein's spectra of three white dwarfs. One has nothing but strong helium

lines; one with very broad hydrogen lines is shown for comparison, together with one which is discussed in subdivision B of this section.

(vii) The carbon stars, in which bands due to carbon molecules dominate the spectrum instead of bands due to oxides, as in *M* and *S* stars, have an abnormally high abundance of carbon relative to oxygen, which will all be used in forming the spectroscopically inaccessible CO molecule. It has therefore not been possible so far to determine the ratio of carbon relative to hydrogen. Bouigue (Bo54a) has found that nearly equal abundances of carbon and oxygen, i.e., a carbon/oxygen ratio between 3 and 6 times normal, could account for the spectroscopic appearance (see Bi54). The carbon/oxygen ratio may vary among the carbon stars (Ke41).

The majority of carbon stars have an apparently normal atmospheric abundance of hydrogen. They may have a higher than normal abundance of the heavy elements (Bi54, Gr54a) (cf. paragraph D (iii) of this section], but no abundance determinations are yet available; technetium has been observed (Me56a). Again, the majority of these stars show bands of $C^{12}C^{13}$ and $C^{13}C^{13}$, indicating that the C^{12}/C^{13} ratio is about 3 or 4 (Mc48), which is near the equilibrium ratio produced in the CN cycle. Thus these stars may be ones in which helium burning ($3He^4 \rightarrow C^{12}$) has been taking place in the core (together, probably, with some neutron production leading to the *s* process). Large-scale mixing may then have set in *before* all the hydrogen in the envelope had been exhausted; the C^{12} passed through a hydrogen-burning shell on its way out to the surface, and the equilibrium ratio of C^{13} was thus produced. As many authors have noted, nitrogen should be abundant in these stars. Two examples of this kind of star, X Cancri (an irregular variable) and HD 54243, will be seen in Plate 2. Such stars at a later stage in their evolutionary history might look like WN stars or the star studied by Münch, HZ 44. However, the increased carbon/oxygen ratio might possibly be due to destruction of oxygen in a sufficiently hot hydrogen-burning region.

(viii) Some rare specimens among the carbon stars have weak or absent CH bands and hydrogen lines, and have been suggested to be deficient in hydrogen (Wu41, Bi53, Bu53). There is no evidence to suggest that in these stars the C^{12}/C^{13} ratio is smaller than that in the solar system ($\sim$90). Most of the group are variables of the R CrB type; RCrB itself was analyzed by Berman (Be35) and found to be very rich in carbon and deficient in hydrogen. There are a few nonvariable examples, two of which are shown in Plate 2. A spectrum of HD 137613 (Sa40, Bi53) has been borrowed from the Mount Wilson files and is shown alongside the two normal carbon stars; absence of C^{13} bands and weakness of H_β are striking. The stars are not a good match in temperature, and hence differences in their line spectra will be seen; however, HD 137613 is hotter than the other two so that H_β would be expected to be stronger in it. Greenstein's spectra of the peculiar carbon star HD 182040, first discussed by Curtiss (Cu16), and the normal carbon star HD 156074, which have similar temperatures, are shown together; the weakness of the CH band and H_γ in the former, while C_2 can be seen, is notable. HD 182040 has no detectable C^{13}.

Such stars are too cool to show helium even if it is in high abundance, although Herbig (He49a) suggested tentative identification of an emission feature at λ 3888.4 in the spectrum of R CrB near minimum light with HeI. Possibly the $3He^4 \rightarrow C^{12}$ reaction has been going on in the core and the products were not mixed to the surface until almost all of the hydrogen in the star had been used up, and no hydrogen-burning shell existed, and hence no C^{13} was produced. The R CrB stars probably belong to Population II. Stars of this sort at a later evolutionary stage might look, for example, like HD 124448, or like those central stars of planetary nebulae which are of WC type, as discussed in (iv) above.

Two stars with a large C^{12}/C^{13} ratio, HD 76396 and HD 112869 (Mc48), have excessively strong CH; they belong to the high-velocity group of CH stars studied by Keenan (Ke42). Another star of the same sort is HD 201626 (No53). If C^{12} has actually mixed out to the surface, then it cannot have gone through a hydrogen-burning region. Nitrogen should then be less abundant than in normal carbon stars. One possibility is that these stars do not possess a hydrogen-burning shell, but have a relatively thin surface region in which hydrogen still exists. Another possibility is that the temperature of an outer hydrogen-burning zone was not high enough. The high-velocity (Population II) character of the CH stars means that they may, at the beginning of their lives, have had a ratio of carbon, nitrogen, and oxygen to hydrogen which was much lower than in Population I stars.

(ix) Miss Roman (Ro52) has found that certain *G*- and *K*-type giants having high velocities have a peculiar appearance in the cyanogen band; she has designated these the "4150" stars. Bidelman (Bi57) suggested that these stars may have a higher than normal carbon abundance. However, he has commented recently that, although these stars have strong cyanogen (CN), they do not show C_2 bands (private communication). Perhaps they are also examples of stars whose surface layers contain material that has recently passed through the CN cycle; the nitrogen/carbon ratio may be larger than normal.

A peculiar *G*-type giant, HD 18474 (Bi53, Hu 57a) and two similar stars studied by Greenstein and Keenan, have weak CH and somewhat weakened CN absorption; these seem to have a low carbon/hydrogen ratio, while the abundance of the metals seems to be normal in two of the stars and low in one which has a high velocity (Gr57a).

To summarize, we conclude that the spectroscopic

indications of the occurrence of hydrogen burning and helium burning in stars show great diversity. Hydrogen burning is demonstrated by complete hydrogen exhaustion (e.g., HD 160641, some white dwarfs); partial hydrogen exhaustion and high nitrogen abundance (e.g., HZ 44, WN stars, υ Sagittarii); occurrence of C^{13} in the equilibrium ratio with C^{12} (e.g., normal carbon stars). Knowledge of the masses of the various stars mentioned here is scanty, but a wide range may be involved, leading to a range in evolutionary tracks and time-scales (see Fig. XI,1). The general evolutionary sequence, however, is probably as follows: (a) the carbon stars (red giants); (b) stars like HD 160641, the WN stars, and HZ 44 ("horizontal branch" stars and their analogy in Population I); (c) the white dwarfs (exhaustion of fuel, end of life).

Evidence for helium burning is less direct. The main problem is that in cool stars the depletion of oxygen during hydrogen burning will always lead to an apparent increase of carbon, because less CO will be formed. Further study of dissociation equilibria of the various molecules, together with a knowledge of the opacity of carbon stars, may throw more light on the problem. However, the WC stars and possibly HD 124448 may have a large increase in carbon; in HZ 44 a smaller increase may have occurred (see Table XI, 2).

The way in which conditions for hydrogen and helium burning depend on the initial mass of the star is at present quite uncertain, and more computed evolutionary tracks, carried through the onset of helium burning, are needed.

B. α Process

A peculiar white dwarf, Ross 640 (Gr56), has strong features attributed to magnesium and calcium. Greenstein's spectrum of this star is reproduced in Plate 1. Although analyses of white dwarf atmospheres, with their extremely large surface gravities, have not been carried out, it seems hard to explain such a spectrum except by the actual presence of large amounts of the α-process elements magnesium and calcium.

One peculiar and perhaps unique white dwarf has no recognizable spectral features except very broad absorptions at $\lambda\lambda$ 3910, 4135, and 4470. Greenstein (Gr56a) has suggested that these may be the helium lines $\lambda\lambda$ 3889 +3965, 4121+4144, and 4438+4472, arising under conditions of extreme pressure, with the obliteration of atomic orbits which would give rise to other lines normally expected. Burbidge and Burbidge (Bu54) have suggested that the features at $\lambda\lambda$ 3910 and 4135 may be due to Si II $\lambda\lambda$ 3854–3863 and 4128–4131, and that the feature at λ 4470 may be due to Mg II 4481, again under very high pressure, and have proposed that the star has a high abundance of the α-process elements magnesium and silicon.

C. Synthesis of Elements in the Iron Peak of the Abundance Curve, and the Aging Effect as It Is Related to this and Other Types of Element Synthesis

Stars which have evolved so far as to have the central temperatures and densities necessary for the equilibrium production of elements in the iron peak become unstable (see Sec. XII). Thus if they spend only a short time in this stage, specimens in which it is occurring or has recently taken place would be expected to be rare (unless stellar remnants—white dwarfs—are left after the explosion). However, if element synthesis has been going on throughout the life-history of our galaxy we might expect to observe an aging effect in that the oldest stars might have the lowest metal abundances. The following observations support this.

(i) Some stars in whose spectra metallic lines are strikingly weak have been observed. Analysis by Chamberlain and Aller (Ch51) of the so-called subdwarfs, HD 19445 and HD 140283, led to an iron abundance 1/10 of normal and a calcium abundance 1/30 of normal for these stars. A recent more detailed analysis of HD 19445 by Aller and Greenstein (Al57b) has confirmed the low iron abundance, which they find to be 1/20 of normal. Plate 2 shows the photographic region of the spectrum of HD 19445 together with the normal *F*7 V star ξ Peg; HD 19445 actually has a slightly lower temperature than ξ Peg (Gr57), and if it were compared with a *G*0 or *G*2 star (where its temperature places it) the difference in the strength of the metallic lines would be even more striking than illustrated in Plate 2. The absolute magnitude of HD 19445, about +5, is very near to the main sequence at the true temperature of the star; certainly the departure from the main sequence is within the error in the trigonometric parallax. The name "subdwarf" is thus seen to be misleading. Most of the "subdwarfs" in this temperature range have been so designated from spectroscopic parallaxes, which are uncertain since the spectral peculiarities of these stars have caused them to be classified too early on low-dispersion spectrograms.

HD 19445 and HD 140283 have high space velocities and belong to Population II. The implication is that they are members of a spherical distribution of stars which condensed before the Galaxy assumed its present flattened form and hence are old stars. Similar extreme weak-line stars, presumably belonging to Population II, e.g., HD 122563 (Gr57), may have low velocities. The fact that calcium is weakened as well as the iron-peak elements means that this aging effect shows, as would be expected, in other element-building processes besides the *e* process. Plate 2 shows that a line of Sr II is also weak (strontium is produced by the *s* process).

Five more stars which have some of the characteris-

tics of Population II were analyzed by Burbidge and Burbidge (Bu56a), by the curve-of-growth method without allowing for the effect of a lowered metal abundance on the opacity. In each star the abundances of a selection among the elements magnesium, aluminum, calcium, scandium, titanium, chromium, manganese, iron, strontium, yttrium, zirconium, and barium were determined. The stars showed a spread in the ratio of the abundances of these elements relative to hydrogen which indicated a range in ages; the most extreme case had a similar ratio to those in HD 19445 and HD 140283. The abundance ratios for strontium, yttrium, and zirconium showed no significant differences from those for iron, but tentative results for barium gave smaller ratios than those for iron (this result was however, uncertain).

(ii) Stars with high space velocities do not belong to the solar neighborhood, and in particular those with high z velocities may be presumed to belong to an older more spherical population. The so-called subdwarfs have very high space motions, but Miss Roman (Ro50, Ro52) and Keenan and Keller (Ke53) found a general correlation between space motion and spectral peculiarities in that high-velocity stars tend to have strong CH and weak CN and metallic lines. A recent analysis of a group of high-velocity stars by Schwarzschild, Schwarzschild, Searle, and Meltzer (Sc57b) has shown that the most extreme example measured by them, ϕ^2 Orionis, has ratios of the metals to hydrogen and of carbon, nitrogen, and oxygen to hydrogen, that are both one quarter of the solar values. The heavy elements also are under-abundant by the same factor as iron. Thus if the different compositions of the so-called subdwarfs, the less extreme high-velocity stars, and the solar-neighborhood stars are due to an aging effect, then elements produced by hydrogen burning, helium burning and the α, e, and s processes are all affected, and in the high-velocity stars the factors are all about the same. Seven high-velocity G-type giants analyzed by Greenstein and Keenan (Gr57a) have values for the metals/hydrogen ratio varying from 0.4 to 0.6 times normal.

(iii) The globular clusters may contain the oldest stars in our galaxy. The age of M3 has been given as 6×10^9 years by Johnson and Sandage (Jo56). It might therefore be expected that the stars in such a cluster would have lower abundances of the metals and other elements than normal. Also, theoretical evolutionary tracks can be made to agree with the observed color-magnitude diagram only if a low metal abundance is used (Ho55). Unfortunately, the brightest main-sequence stars of this cluster (in which turbulence and opacity should have only small effects on abundance determinations) are too faint for spectrophotometric analysis. However, information may be derivable from the colors. The two-color (U−B, B−V) plot for M3 lies well above the normal solar-neighborhood plot throughout its length, while the so-called subdwarf HD 19445 [see C (i)] lies very near to the M3 plot. This position of HD 19445 may be the result of different "blanketing" of the continuous spectrum by the spectrum lines. All lines are weakened in HD 19445; the metallic lines are more numerous in the blue (B) than in the visual (V) spectral region, and still more numerous in the ultraviolet (U). The effect of weakened metallic lines would therefore be to increase the blue intensity relative to the visual, and to increase the ultraviolet relative to the blue. Qualitatively therefore, as discussed by Johnson and Sandage (Jo56), this is in the same sense as that required to explain the abnormal position of HD 19445 on the U−B, B−V diagram. An accurate evaluation of the effect is currently being made by Sandage, Burbidge, and Burbidge. If this explanation is found to be satisfactory then we shall have indirect proof that the stars of M3 have low metal abundances. Sandage has shown by an approximate calculation that if the amount of light subtracted from the sun's continuum by the spectrum lines is added in the appropriate regions (without allowing for any redistribution of energy caused by blanketing), the effect is to shift the sun on to the M3 line in the (U−B, B−V) plot, at a value of B−V corresponding to a normal $F5$ star.

(iv) Studies of colors and spectra of external galaxies are still too uncorrelated for us to make deductions concerning abundance variations. It may, however, be worth entertaining the possibility that some of the elliptical galaxies which have much earlier-type spectra than would correspond to their color classes may prove to have low calcium and iron abundances, relative to hydrogen.

(v) Apart from the foregoing discussion of possible aging effects (in other processes besides the e process), no undoubted examples of stars exhibiting the e-process abundances predominantly are known. However, a speculative possibility may be considered. The white dwarf van Maanen 2 shows strong, very broadened lines of iron (Gr56); possibly it has a high abundance of iron. Such stars may be the remnants (perhaps after large mass loss) of stars which reached a central temperature sufficient to build the iron peak, but did not suffer complete catastrophic explosion.

D. *s* Process

Three groups of giant stars have abundance anomalies that indicate the occurrence of the s process in their interiors. These are (i) the S stars; (ii) the Ba II stars; and (iii) the carbon stars.

(i) The S stars probably have luminosities in the range of normal giants. Elements with neutron magic-numbers appear more predominantly in the spectra of S stars than in M stars of the same temperature (Me47, Al51, Bu52, Bi53a, Me56, Te56). It has been realized for some time that this must be a real atmospheric abundance effect and not an ionization,

TABLE XI,3. Overabundance ratios for the sum of the s-process isotopes of each element in HD 46407, calculated on the assumption that the observed overabundance ratios are due solely to building of the s-process isotopes.

Element	Observed ratio y	Suess and Urey abundances s	$r+p$	Calculated ratio for s-isotopes y'	$\langle\sigma N\rangle y'$
Strontium	4.7	18.79	0.106	4.7	381
Yttrium	7.8	8.9	···	7.8	1326
Zirconium	4.8	52.92	1.53	4.9	884
Niobium	5.2	1.00	···	5.2	832
Molybdenum	3.2	1.4185	1.0015	4.8	343
Ruthenium	7.8	0.6445	0.8452	16.7	1837
Barium	14	3.4326	0.2253	14.9	281
Lanthanum	9.8	2.00	0.0018	9.8	647
Cerium	9.6	2.00	0.2601	10.7	407
Praseodymium	28	0.40	···	28	672
Neodymium	14	1.276	0.1630	15.7	443
Samarium	13	0.3080	0.3448	26.4	4050
Gadolinium	3.0	0.0147	0.668	93	3255
Ytterbium	2.0	0.12346	0.0944	2.8	78
Tungsten	9	0.326	0.1646	13.0	914

TABLE XI,4. Search for r-process elements in HD 46407.

Element	Whether observed in the sun	Sensitive lines: whether accessible	Whether strengthened in HD 46407
Selenium	No	No	No evidence
Bromine	No	No	No evidence
Krypton	No	No	No evidence
Tellurium	No	No	No evidence
Iodine	No	No	No evidence
Xenon	No	No	No evidence
Osmium	Yes	Yes	Inconclusive[a]
Iridium	Yes	Yes	Probably no[b]
Platinum	Yes	Yes	Very probably no[c]
Gold	Yes	Yes	No evidence[d]

[a] All lines in the observed wavelength range of HD 46407 are too badly blended for any identifications to be made.
[b] At the wavelengths of the strongest lines in the observed range, *no lines* are visible, either in HD 46407 or in the standard star.
[c] There are four possible identifications, for all of which the lines in HD 46407 are equal to those in the standard star.
[d] Sensitive line is outside the observed range in HD 46407. Other lines are not seen.

dissociation, or excitation effect. Fujita (Fu39, Fu40, Fu 41) suggested that the S stars have carbon abundances intermediate between M and the carbon stars. The appearance of strong lines of Tc I in S stars (Me52) has been previously described as a good indicator of current nuclear activity and mixing on a time scale $\sim 10^5$ years in these stars. In Plate 3 we show two portions of the spectrum of the S-type long-period variable star R Andromedae obtained by Merrill, together with a standard $M5$ III star; this spectrum was taken from the plate files at Mount Wilson Observatory. The first portion shows the lines of Tc I in R Andromedae; the second shows a region where ZrO bands are present in R Andromedae, and TiO bands are seen in the M-type star. The strength of Ba II λ 4554 in R Andromedae will also be noted. The two stars are not a perfect match in temperature, so that some differences in the line spectra will also be noted. R Andromedae is somewhat cooler than the M-type star, and since it is a long-period variable, emission lines will be seen in its spectrum, notably at H_γ.

(ii) From the spectroscopic point of view the rarer Ba II stars (Bi51) represent an easier problem to analyze, since their temperatures are considerably higher, their spectra being of types G-K, while their spectral peculiarities link them closely with the S stars. They have a greater-than-normal abundance of carbon, probably intermediate between the M and the carbon stars. A recent curve-of-growth analysis of HD 46407, a typical Ba II star (Bu57a), has yielded abundances of a large number of elements relative to those in the standard $G8$ III star κ Geminorum, which is very similar in temperature and luminosity and was used as a comparison. Portions of the spectra of HD 46407 and κ Geminorum are shown in Plate 3; the most prominent of the features that are strengthened in the Ba II star are marked. Abundances of all elements which could be studied, lighter than $A=75$, and in particular elements in the iron peak, are normal. Most of the elements with $A>75$, which are spectroscopically observable, are overabundant; this includes in particular those with neutron magic-numbers. The observed abundance ratios, relative to normal, are given in column 2 of Table XI,3. With the possible exception of gadolinium, all these elements have a large proportion of their isotopes produced in the s process, as is shown in the appendix. On the other hand, with the possible exception of ytterbium, all of the elements with $A>75$ which we have assigned mainly to the s process, and which are spectroscopically observable, are found to be overabundant.

The observations in HD 46407 refer only to the sum of all isotopes of each element, and we wish to derive the true overabundances of those isotopes actually built by the s process. From the appendix we have obtained the sum of the solar-system abundances of the isotopes of each element in Table XI,3 which are built by the s process, and by the r and p processes together. If these sums are denoted by s and $r+p$, respectively, and if the observed and true overabundances are y and y', respectively, then we have

$$y(s+r+p)=y's+r+p.$$

Values of s, $r+p$, and y' are given in columns 3, 4, and 5 of Table XI,3.

Finally, we have taken the mean vale of σN for all the s-process isotopes of each element, and the product of this and y' is given in the last column. Under conditions of steady flow this product should be constant (see Sec. VI), and this is the case to within a factor of about 2, if we exclude ruthenium, samarium, gadolinium, ytterbium. Of these, only samarium has a well-determined abundance, and the possibility that cross sections for this element in the appendix are too large was discussed in Sec. VI. The estimates of σ for Gd^{154}, Gd^{155}, and Yb^{89} may also be too large, while

that for Ba^{134} may be too low. If further abundance determinations confirm the low value of y for ytterbium then it may be found that a smaller proportion of its abundance should be assigned to the s process.

Europium, both of whose isotopes have been assigned to the r process, turned out not to be overabundant in HD 46407. A search for the most abundant of the r process elements in the spectrum of HD 46407 was inconclusive; most of these elements cannot be identified in HD 46407 or in the standard star. However, none were found to be strengthened in HD 46407. The results of this search are given in Table XI,4. This table is given in order to show how difficult it is to identify spectrum lines due to many of the r-process heavy elements, let alone determine abundances, in stars other than the sun.

The elements copper, zinc, gallium, and germanium are all produced predominantly by the s process, yet the first two and probably the last are not overabundant in HD 46407 (we have no evidence concerning gallium). These elements fall on the steep part of σN *versus* A plot in Fig. VI,3, and this suggests, therefore, that there is a plentiful supply of neutrons produced in Ba II stars, giving the shorter of the two capture-time-scales discussed in Sec. II, and leading to building of elements according to the flat part of the curve in Fig. VI,3. If this process occurs in a volume containing $10^{-2}-10^{-3}$ of the mass of the star, then subsequent mixing could produce the observed overabundances; normal abundances would be expected to be observed for copper, zinc, gallium, and germanium, and depletion of the iron-peak elements would not be detectable. Such a copious flux of neutrons means that interior temperatures of $\sim 10^8$ degrees must have been reached. However, we have independent indications, from the presence of an excess amount of carbon in this and other Ba II stars (as also in the S stars), that a temperature sufficient for the reaction $3He^4 \rightarrow C^{12}$ to occur has been achieved.

Tc I lines have not been observed in Ba II stars; before it can certainly be said that technetium is absent a search should be made for Tc II in the ultraviolet. Possibly the Ba II stars represent a later evolutionary stage than the S stars, or they may evolve from stars in a different mass range in which mixing to the surface takes a time $\gg 10^5$ years.

HD 26 is a Ba II star with Population II characteristics (abnormally strong CH, weak metallic lines, and a high velocity). The strongest spectrum lines include those of the rare earths [Vrabec, quoted by Greenstein (Gr54a)]. This suggests that the s process is also taking place in the oldest stars.

(iii) As was mentioned in paragraph A(vii), the carbon stars also have apparent overabundances of the heavy elements, less striking than in the S stars, together with the presence of technetium. Actual abundances have not been determined, but it seems as though the s process has been occurring here also, perhaps in a smaller volume of the star, while a core containing more carbon is developed. More abundance determinations, together with computations of evolutionary tracks of stars, will be helpful in elucidating the evolutionary differences between M, C, S, and Ba II stars.

E. r Process

The outstanding piece of observational evidence that this takes place is given by the explanation of the light curves of supernovae of Type I as being due to the decay of Cf^{254} (Bu56, Ba56), together with some other isotopes produced in the r process. Further evidence can be obtained only by interpreting the spectra of Type I supernovae, a problem which has so far remained unsolved.

F. p Process

No astrophysical evidence is at present available for the occurrence of this process. The present theory might suggest that isotopes built in this way should be seen in the spectra of supernovae of Type II.

G. x Process

The situation as far as the astrophysical circumstances of the x process are concerned has been discussed in Sec. X.

H. Nuclear Reactions and Element Synthesis in the Surfaces of Stars

The peculiar A stars, in which large magnetic fields have recently been found (see Ba57), have long been known to have anomalously strong (and often variable) lines of certain elements in their spectra. Abundances of all the observable elements have been determined in two magnetic stars, α^2 Canum Venaticorum and HD 133029, and in one peculiar A star, HD 151199, in which a magnetic field has not been detected, owing to the greater intrinsic width of the spectrum lines, but may well be present (Bu55, Bu55a, Bu56b). The results are given in Table XI,5, which should be compared with Table XI,3 and the discussion of the Ba II star HD 46407. At first glance the results are very similar, particularly in the fact that predominantly the elements with neutron magic-numbers are overabundant, but there are some striking differences, as follows: (a) barium is overabundant in HD 46407, and normal in the peculiar A stars; (b) europium is the most prominently overabundant element in the peculiar A stars, and is normal in HD 46407; (c) silicon, calcium, chromium, and manganese are normal in HD 46507 while calcium is underabundant and the rest are overabundant in the peculiar A stars; and (d) the overabundance ratios of the strontium group and the rare earth group differ by factors of about 2 in HD 46407 and by factors varying from 2 in HD 151199 to 40 in α^2 Canum Venaticorum.

TABLE XI,5. Abundances of the elements in three "peculiar A" stars, relative to those in a normal star.

Element	Magic number isotope (A,N) for anomalous elements	α^2CVn	HD 133029	HD 151199
Magnesium		0.4	1.4	1.2
Aluminum		1.1	2.2	
Silicon	28, 14	10	25	1.3
Calcium	40, 20	0.02	0.05	2.6
Scandium		0.7		
Titanium		2.6	2.3	
Vanadium		1.3	3.2	
Chromium	52, 28	5.2	10.3	1.8
Manganese		16	15:	9
Iron		2.9	4.1	1.1
Nickel		3.0	2	
Strontium	88, 50	14	11.5	65
Yttrium	89, 50	20:		
Zirconium	90, 50	30	40	
Barium	138, 82	$\leq$0.9		0.6
Lanthanum	139, 82	1020:	200:	
Cerium	140, 82	400	190	
Praseodymium	141, 82	1070	630	
Neodymium	142, 82	250	200	
Samarium	144, 82	410	260	
Europium		1910	640	130
Gadolinium		810	340	
Dysprosium		760	460	
Lead	208, 126	1500:	1500:	

Clearly, some kind of element synthesis has been taking place in the peculiar A stars (Fo55a, Bu57). Since it is unlikely that the products of any nuclear reaction in the interior could have mixed to the surface, and since an energy source is available in the atmosphere in the form of the magnetic fields, we have suggested that this represents a special process of element synthesis which can only take place in stellar atmospheres where there is acceleration of nucleons by electromagnetic activity to energies far above thermal. The predominance of magic-number nuclei among the overabundant elements indicates that the type of synthesis probably involves neutrons; in fact, it may be essentially an s process taking place in a proton atmosphere. However, the buildup may be by deuteron stripping reactions (d,p) as well as or instead of by (n,γ) processes; (α,n), $(\alpha,2n)$, and $(\alpha,3n)$ may also occur. Free neutrons are thought to be produced by (p,n) processes on the light nuclei, and their presence in a hydrogen atmosphere means that equilibrium abundances of both deuterium and He^3 will be produced. Evidence for the presence of deuterium in magnetic stars is at present lacking; evidence for the possible presence of He^3 in one of them, 21 Aquilae, has been discussed by Burbidge and Burbidge (Bu56c).

The activity in magnetic stars is thought to take place in "spot" regions, with a duration of only a few seconds (the time-scale is limited by the rate of energy loss by bremsstrahlung), but must continue over long time-scales comparable with the lifetime of the stars in order to build the observed anomalies of the heavy elements, so that each gram of the surface layer is processed many times. An important condition for the detection of the anomalies is that no mixing between the surface layer and the interior should occur, or it would dilute the products of nuclear activity so that they would not be detectable. This condition happens to be fulfilled in the A-type stars, where the convective layer extends from the photosphere inwards to a depth of only 1000 km (cf. Fo55a).

XII. GENERAL ASTROPHYSICS

A. Ejection of Material from Stars and the Enrichment of the Galaxy in Heavy Elements

Table XII,1 gives a tentative assessment of the production of elements in the Galaxy, the table being subdivided in accordance with the various processes described in the previous sections.

The values given in the third column were obtained in the following way. The fractions by weight of the various groups of elements in the solar system have already been given in Table II,1. These fractions were reduced by a factor 2 (except for hydrogen) to make allowance for the lower concentrations of the heavy elements, relative to the sun, in the stars which are believed to comprise the main mass of the Galaxy. It is clear that some such reduction must be made, and the factor 2, although somewhat uncertain, is based on the work of Schwarzschild *et al.* (Sc57b; cf. also earlier work mentioned therein). The resulting fractions, when taken together with a total mass of 7×10^{10} $M_\odot$ for the Galaxy (Sc56), yield the values shown in the third column of Table XII,1.

The products of the synthesizing processes must be distributed in space. Emission from stellar atmospheres apart, two ejection mechanisms are mentioned in the fourth column of Table XII,1, supernovae, and red giants and supergiants. These are believed to be of main importance, but matter is also ejected from novae, close binary systems (St46, Wo50, St57a), P Cygni and WR stars, and planetary nebulae (cf. Gu54, St57). Leaving aside the x process, the association between the second and fourth columns of Table XII,1 was made on the basis that only in supernovae are temperatures of thousands of millions of degrees reached. Reasons for this view have been given in Sec. III,F. In contrast, temperatures of hundreds of millions of degrees may be expected in red giants and supergiants (Sec. V, B), and in the other contributors, novae, binaries, etc. Thus, processes of synthesis requiring temperatures of order 10^8 degrees have been associated with the giants, whereas processes requiring temperatures of order 10^9 degrees have been associated with supernovae.

Estimates of the total amount of material which may have been distributed in space by supernovae and red giants and supergiants, assuming a constant

TABLE XII,1.

Elements	Mode of production	Total mass in galaxy ($M_\odot$ as unit)	Astrophysical origin	Total mass of all material ejected over lifetime of galaxy ($M_\odot$ as unit)	Required efficiency
He	H burning	8.1×10^9	Emission from red giants and supergiants	2×10^{10}	0.4
D	x process?	7.5×10^6?	Stellar atmospheres? Supernovae?	?	?
Li, Be, B	x process	8.5×10^2	Stellar atmospheres	?	?
C, O, Ne	He burning	4.3×10^8	Red giants and supergiants	2×10^{10}	2×10^{-2}
Silicon group	α process	4.0×10^7	Pre-Supernovae	2×10^8	0.2
Silicon group	s process	8.5×10^6	Red giants and supergiants	2×10^{10}	4×10^{-4}
Iron group	e process	2.4×10^7	Supernovae	2×10^8	0.1
$A>63$	s process	4.5×10^4	Red giants and supergiants	2×10^{10}	2×10^{-6}
$A<75$	r process	5×10^4	Supernovae Type II	1.7×10^8	3×10^{-4}
$A>75$	r process	10^4	Supernovae Type I	3×10^7	3×10^{-4}
$A>63$	p process	1.3×10^2	Supernovae Type II	1.7×10^8	10^{-6}

rate of star formation and death during the lifetime of the Galaxy, are given in the fifth column of Table XII,1 and were made in the following way.

The rates of supernovae were taken to be 1 per 300 years for Type I and 1 per 50 years for Type II, as estimated by Baade and Minkowski (Ba57a), and the age of the Galaxy was taken as 6×10^9 years. The average mass emission per supernova is not easy to estimate (see Sec. XII,C), but we have taken it to be 1.4 $M_\odot$. The exponential-decay light curve is characteristic of Type I and not of Type II supernovae; this has been taken to indicate that the rapid production and capture of neutrons, leading to synthesis of the heaviest r-process elements, occurs only in Type I supernovae. The lighter r-process elements are probably made in Type II supernovae, while the products of the α and e processes may perhaps be made in either type.

The estimate for the ejection of mass by red giants and supergiants was based on the assumption that such stars shed most of their material into space during the giant phases of their evolution. Thus Deutsch (De56) has shown that matter is streaming rapidly outwards from the surface of an M-type supergiant, and he considers from the spectroscopic data that probably all late-type supergiants are also ejecting material at a comparably rapid rate. There is also some evidence that normal giant stars are ejecting matter, although probably on a smaller scale. Hoyle (Ho56b) has indicated that mass loss may have a more important effect on the evolution of giants and supergiants of very low surface gravity than have nuclear processes. With these considerations in mind, it is necessary to estimate the fraction of the mass of the Galaxy that has been condensed into stars which have gone through their whole evolutionary path.

The luminosity and mass functions for solar-neighborhood stars have been studied by Salpeter (Sa55a). He has shown that stars which lie on the main sequence now with absolute visual magnitudes fainter than +3.5 (which form the bulk of stars at present in existence) make up 55% of the total mass. Stars which formed with magnitudes brighter than +3.5 (the majority of which have gone through their whole evolutionary path) make up the remaining 45%. We find from his mass function that the average mass of the stars brighter than +3.5 was 4 $M_\odot$. Since the upper limit to the mass of a white dwarf is about 1.4 $M_\odot$, we find that, of the remaining 45%, 16% may lie in white dwarfs and 29% in the form of ejected gas. Thus the division of mass between stars which have not evolved, white dwarfs, and material which has been processed by stars is approximately 4:1:2. Extrapolating these results to the Galaxy as a whole, we find that 2×10^{10} $M_\odot$ has been processed by stars. Some of this is included in the estimate of matter ejected by supernovae, but the majority will have been processed at temperatures $\lesssim10^8$ degrees, and ejected by red giants and supergiants.

The efficiency of production of each group of elements in the total ejected material that is necessary to explain the abundances in column 3 of Table XII,1 is obtained by dividing column 3 by column 5, and is given in the last column. It must be emphasized here that the estimates given in Table XII,1 are very preliminary. Even on an optimistic view they might well prove incorrect by as much as a factor 3. It is noteworthy, however, that the emission seems adequate to explain the synthesis requirements, except possibly in the case of deuterium, and helium, which is the only element needing an efficiency near to unity for its production. It will be clear from the foregoing that our estimates are not accurate enough to establish whether or not it is necessary to assume some helium in the original matter of the Galaxy.

The efficiency necessary to produce the r-process elements is considerably smaller than that required to give the observed supernova light curves (see Sec. XII,C). It should, in addition, be reiterated that we have no evidence at all concerning the abundances of the r-only and p-process isotopes outside the solar system, as we pointed out in Sec. XI.

If the universal value of the deuterium to hydrogen ratio turns out to be of order 10^{-4}, it will be difficult to provide for an adequate abundance of deuterium by processes of an electromagnetic nature occurring in stellar atmospheres, as has already been pointed out in Sec. X,C. The possible alternative process of origin, also discussed in Sec. X,C, rests on considerations to be given in Sec. XII,B. At sufficiently high temperatures ($T>7\times10^9$ degrees) elements of the iron group break down into α particles and neutrons. For example, $Fe^{56}\rightarrow13\alpha+4n$. In Sec. XII,B the onset of a supernova is associated with such breakdown reactions developing in the central regions of a star. In Type II supernovae, hydrogen is present in the outer expanding envelope, and capture of the neutrons by relatively cool hydrogen yields deuterium without subsequent disintegration. As was pointed out in Sec. X,C, if one supernova converted one solar mass of its envelope into deuterium and was then diluted by mixture with 10^4 solar masses, the terrestrial abundance would result. These estimates of production and dilution are not at all unreasonable. On this basis the galactic abundance of deuterium would not necessarily be as large as in the solar system.

The balance between the observed solar-system abundances and the material ejected from stars has been made on the assumption of a constant rate of star formation and death during 6×10^9 years. There are some considerations which suggest that the rate may have been greater in the early history of the Galaxy.

First, recent work on the 21-cm radiation by hydrogen, reviewed by van de Hulst (Hu56), has shown that neutral atomic hydrogen makes up only a few percent of the total mass of both our Galaxy and M31 (and it probably comprises the major part of the gas and dust). Star formation is currently occurring only in Population I regions, which comprise about 10% of the mass of the Galaxy (by analogy with the ratio found by Schwarzschild (Sc54) in M31). Originally the Galaxy must have been composed entirely of gas. Thus it would seem probable that star formation and element synthesis would have occurred at a greater rate early in the life of the Galaxy, and, as a larger and larger fraction of the mass became tied up in small stars and in white dwarfs, so nuclear activity would have been a decreasing function of the age of the Galaxy.

Second, the solar system is 4.5×10^9 years old, and the uranium isotope ratio (Sec. VIII) gives 6.6×10^9 years as the minimum age of the r-process isotopes. Thus the abundances that we are discussing (which are half the solar system abundances) must have been formed (by all 8 synthesizing processes) in a time less than the age of the Galaxy, say $1\text{–}2\times10^9$ years. Also, there are apparently large abundance differences between extreme Population II stars such as the globular clusters M92 (and M3) and, for example, the old galactic cluster M67, while the ages that have been given do not differ correspondingly and are both $\sim6\times10^9$ years (Jo56, Sa57a). It is of interest to mention recent investigations of the color-magnitude diagram of the nearby stars by Sandage (Sa57b) and by Eggen (Eg57), based on the data of Eggen (Eg55, Eg56) and the Yale Parallax Catalogue (Ya52). There are a few stars to the right of the main sequence at $M=3.5$ to 4; if the parallaxes are accurate then they may be either stars which are still contracting and have not yet reached the main sequence, or they may have evolved off the main sequence. In the latter case, their ages would be $\gtrsim8\times10^9$ years.

These facts may be reconciled if we suppose that star formation and death in the originally spherical gaseous Galaxy occurred at a sufficient rate to produce element abundances which then recondensed into the oldest star systems that we see today in the halo. The initial differentiation of Populations I and II was interpreted as one of location in the Galaxy and of comparative youth (spiral arms) and age (halo, nucleus). However, it is clear that a number of subclassifications are demanded (Pa50, We51, Bu56a). While the oldest halo stars that we see today were condensing, we may suppose that concentration towards the center and the equatorial plane was occurring, and the formation of stars was favored by conditions in the flattening higher-density region, leading to a rapid gradient of heavy-element enrichment so that star systems forming in the disk at nearly the same time as the globular clusters had higher abundances of heavy elements. Only gas will contract in this way; stars, once formed, become "frozen" in the shape of the system at the time of their formation. The remaining gas became concentrated into the spiral arms, so that these were the only places where any star formation could any longer occur. If, during the first $1\text{–}2\times10^9$ years in the life of the Galaxy, the average proportion of the mass in which star formation was occurring is taken as 50% (instead of 10% as now), then the increased rate of nuclear activity would compensate for the shorter time scale for element synthesis.

Finally, it is of some importance to discuss the cosmological background to this problem. In an evolutionary explosive cosmology we have that, if T is the age of a system, $T_{\text{universe}}>T_{\text{galaxy}}>T_{\text{some clusters}}>T_{\text{sun}}$. Apart from the products of an initial ylem, if this type of initial condition is proposed, the Galaxy will condense out of primeval hydrogen, and the synthesis will go on as we have already described it. It is of interest that possibly the ylem production of deuterium could be incorporated in a model of this sort. Alternatively, in a steady-state cosmology we have only that $T_{\text{galaxy}}>T_{\text{some clusters}}>T_{\text{sun}}$, and we should expect that the initial condensation of the Galaxy would take place out of gas which already had a weak admixture of the heavy elements that had been synthesized in other galaxies. Thus it might be that a search for low-mass stars of pure hydrogen would present

a possible cosmological test, although such stars might take so long to condense (He55) that they would have become contaminated with the products of the death of massive stars before they had been able to contract themselves into self-contained stars.

It is also of interest to remark on the final stages of galactic evolution. Although, as remarked in the Introduction, the lowest energy state of the elements would be reached when the galaxy was transmuted to iron, it appears that the astrophysical circumstances lead to the matter being finally all contained in white dwarfs (degenerate matter), so that such an extreme condition as an iron galaxy may never be reached.

It should be emphasized that the present value of $1/H$ determined from the red-shift measurements by Humason, Mayall, and Sandage (Hu56a) is 5.4×10^9 years, and this in the evolutionary cosmologies is a time of the order of the age of the universe. However, this value is in conflict with the ages of some star clusters in our galaxy. Thus any attempt to relate the synthesis of the elements in the galaxy through a number of necessarily complex steps to a particular cosmological model must await the resolution of this dilemma.

B. Supernova Outbursts

In the following discussion we consider the course of evolution that in our view leads to the outbursts of supernovae. According to a well-known calculation by Chandrasekhar the pressure balance in a star cannot be wholly maintained by degeneracy for masses greater than a certain critical mass. For pure He^4 this critical mass turns out to be 1.44 $M_\odot$, while for pure iron the value is 1.24 $M_\odot$.

It follows that stars with masses greater than the critical value cannot be in mechanical support unless there is an appreciably temperature contribution to their internal pressures. Mechanical support therefore demands high internal temperatures in such stars.

Our arguments depend on these considerations. We are concerned with stars above the Chandrasekhar limit, and assume that mechanical support is initially operative to a high degree of approximation. This is not a restriction on the discussion, since our eventual aim will be to show that mechanical support ceases to be operative. A discussion of catastrophic stars would indeed be trivial if a lack of mechanical support were assumed from the outset.

The next step is to realize that a star (with mass greater than the critical value) must go on shrinking indefinitely unless there is some process by which it can eject material into space. The argument for this startling conclusion is very simple. Because of the high internal temperature, energy leaks outwards from the interior to the surface of the star, whence it is radiated into space. This loss of energy can be made good either by a slow shrinkage of the whole mass of the star (the shrinkage being "slow" means that mechanical support is still operative to a high degree of approximation), or by a corresponding gain of energy from nuclear processes. But no nuclear fuel can last indefinitely, so that a balance between nuclear energy generation and the loss from the surface of the star can only be temporary. For stars of small mass the permissible period of balance exceeds the present age of the galaxy, but this is not so for the stars of larger mass now under consideration. Hence for these stars shrinkage must occur, a shrinkage that is interrupted, but only temporarily, whenever some nuclear fuel happens for a time to make good the steady outflow of energy into space.

Shrinkage implies a rising internal temperature, since mechanical support demands an increasing thermal pressure as shrinkage proceeds. It follows that the internal temperature must continue to rise so long as the critical mass is exceeded, and so long as mechanical support is maintained. The nuclear processes consequent on the rise of temperature are, first, production of the α-particle nuclei at temperatures from $1\text{–}3\times10^9$ degrees and, second, production of the nuclei of the iron peak at temperatures in excess of 3×10^9 degrees. It is important in this connection that the temperature is not uniform inside a star. Thus near the center, where the temperature is highest, nuclei of the iron peak may be formed, while outside the immediate central regions would be the α-particle nuclei, formed at a somewhat lower temperature, and still further outwards would be the light nuclei together with the products of the s process, formed at a much lower temperature. Indeed some hydrogen may still be present in the outermost regions of the star near the surface. The operation of the r process turns out to depend on such hydrogen being present in low concentration. The p process depends, on the other hand, on the outer hydrogen being present in high concentration.

Turning now to the nuclei of the iron peak, the statistical equations given in Sec. IV show that the peak is very narrow at $T\sim3\times10^9$ degrees, the calculated abundances falling away sharply as the atomic weight either decreases or increases from 56 by a few units. At higher temperatures the peak becomes somewhat wider, ranging to copper on the upper side and down to vanadium on the lower side. Beyond these limits the abundances still fall rapidly away to negligible values with one exception, the case of He^4. From (6) in Sec. IV we find that the statistical equations yield the following relation between the abundances of He^4 and Fe^{56}

$$\log n(4,2)=34.67+\tfrac{3}{2}\log T_9+\frac{1}{14}\{\log n(56,26)-36.40-\tfrac{3}{2}\log T_9\}-\frac{34.62}{T_9}+\frac{\theta}{7}$$

or

$$\log n(4,2)-\frac{\log n(56,26)}{14}=32.08+1.39\log T_9-\frac{34.62}{T_9}+\frac{\theta}{7}.$$

Under the conditions of the present discussion the densities of free protons and neutrons are comparable and small compared with the helium and iron densities, so that the θ term is small and is neglected.

We consider the application of this equation for a fixed value of the total density equal to 10^8 g/cm^3. This is a plausible value for $T \simeq 7\times10^9$ degrees ($T_9 \simeq 7$), and is the mean density of a stellar core of $1M_\odot$ with a radius of 1.7×10^8 cm. Supposing that conditions in this core are such that the mass is equally divided between He^4 and Fe^{56}, then log $(4,2)=30.88$ and log $n(56,26)=29.73$, since no other nucleus makes an appreciable contribution to the density. Substituting in the equation above we find that then $T=7.6\times10^9$ degrees. Thus when T rises above about 7×10^9 degrees there is a very considerable conversion of iron into helium.

This result is not very dependent on the particular value of 10^8 g/cm^3 used above, since the critical temperature at which the masses of helium and iron are comparable increases only slowly with the density (Ho46). The density cannot differ very much from 10^8 g/cm^3. Thus a value much less than 10^6 g/cm^3 would be most implausible for $T\sim7\times10^9$ degrees, while a value appreciably greater than 10^8 g/cm^3 is forbidden by the condition that there must be an appreciable thermal contribution to the pressure.

The argument now runs as follows. Let us suppose that the temperature increases to 8.2×10^9 degrees. In this case the right-hand side of the equation increases by 0.4 and in order that the equation is satisfied, the difference between $\log n(4,2)$ and $(1/14)\log n(56,26)$ must also increase by 0.4. However, $\log n(4,2)$ can only increase by 0.3 since at this point the total mass of the core will be in the form of helium. Thus an increase in temperature of 0.6×10^9 degrees implies that the numerical value of the term $(1/14)\log n(56,26)$ is decreased by at least 0.1, and this implies a drop in the amount of iron of at least a factor of 25. So we conclude that an increase in temperature from 7.6 to 8.2×10^9 degrees causes nearly all of the iron to be converted to helium; i.e., from the situation that the mass is equally divided between helium and iron we reach conditions in which $\sim98\%$ of the mass is helium and $\sim2\%$ is iron. A similar conclusion follows for other values of the total density. However, for much higher temperatures the term in θ becomes important, reflecting the approach toward a neutron core which would be achieved at extremes of density and temperature.

To convert 1 g of iron into helium demands an energy supply of 1.65×10^{18} ergs. This may be compared with the total thermal energy of 1 gram of material at 8×10^9 degrees which amounts to only 3×10^{17} ergs. Evidently the conversion of iron into helium demands a supply of energy much greater than the thermal content of the material. The supply must come from gravitation, from a shrinkage of the star, and clearly the shrinkage must be very considerable in order that sufficient energy becomes available. This whole energy supply must go into the conversion and hence into nuclear energy, so that very little energy is available to increase the thermal content of the material. However, instantaneous mechanical stability in such a contraction would demand a very large increase in the internal temperature. Thus we conclude that in this contraction, in which the thermal energy is, by hypothesis, only increased by a few percent, there is no mechanical stability, so that the contraction takes place by free fall inward of the central parts of the star. At a density of 10^8 g/cm^3, this implies an implosion of the central regions in a time of the order of 1/5 of a second ($\tau \simeq (4/3\pi G\rho)^{-\frac{1}{2}}$ and in our view it is just this catastrophic implosion that triggers the outburst of a supernova.

Before considering the consequences of implosion, it is worthwhile emphasizing that loss of energy by neutrino emission, the urca process (Ga41), is very weak compared to the process described above. Neutrino emission does not rob a star of energy at a sufficiently rapid rate to demand catastrophic implosion, though at temperatures in excess of $\sim3\times10^9$ degrees it does promote a much more rapid shrinkage of the star than would otherwise arise from the loss of energy through normal radiation into space. The reaction $Mn^{56}(\beta^-\nu)Fe^{56}$ with a half-life of 2.58 hr probably gives the main contribution to the loss of energy through neutrino emission. The binding energy of Mn^{56} is less than that of Fe^{56} by 2.895 Mev. Thus, under conditions of statistical equilibrium near $T=5\times10^9$ degrees, this implies that Mn^{56} is less abundant than Fe^{56} by a factor $\sim10^3$. Hence, it is reasonable to suppose that a fraction of about 10^{-3} of the mass is in the form of Mn^{56}. The beta decay of this nucleus leads to an energy loss of about 3×10^{12} ergs per gram of Mn^{56}. Thus a time-scale of about 10^8 sec is demanded to produce an energy loss by neutrino emission comparable with the thermal energy possessed by the stellar material. Hence we conclude that under these conditions in the stellar core, the urca process is quite ineffective as compared with the refrigerating action of the conversion of Fe^{56} to He^4.

The last question to be discussed is the relationship between the *implosion* of the central regions of a star and the *explosion* of the outer regions. Two factors contribute to produce explosion in the outer regions. The temperature in the outer regions is very much lower than the central temperature. Because of this the outer material does not experience the same extensive nuclear evolution that the central material does. Particularly, the outer material retains elements that are capable of giving a large energy yield if they become subject to sudden heating, e.g., C^{12}, O^{16}, Ne^{20}, Ne^{21}, He^4, and perhaps even hydrogen. The second point concerns the possibility of the outer material experiencing a sudden heating. Because under normal

conditions the surface temperature of a star is much smaller than the central temperature, material in the outer regions normally possesses a thermal energy per unit mass that is small compared with the gravitational potential energy per unit mass. Hence any abnormal process that causes the thermal energy suddenly to become comparable with the gravitational energy must lead to a sudden heating of the outer material. This is precisely the effect of an implosion of the central regions of a star. Consequent on implosion there is a large-scale conversion of gravitational energy into dynamical and thermal energy in the outer zones of the star.

One last point remains. Will the gravitational energy thus released be sufficient to trigger a thermonuclear explosion in the outer parts of the star? The answer plainly depends on the value of the gravitational potential. Explosion must occur if the gravitational potential is large enough. In a former paper (Bu56) it was estimated that a sudden heating to 10^8 degrees would be sufficient to trigger an explosion. This corresponds to a thermal energy $\sim 10^{16}$ ergs per gram. If the gravitational potential per gram at the surface of an imploding star is appreciably greater than this value, explosion is almost certain to take place. For a star of mass 1.5 $M_\odot$, for instance, the gravitational potential energy per unit mass appreciably exceeds 10^{16} ergs per gram at the surface if the radius of the star is less than 10^{10} cm. At the highly advanced evolutionary state at present under consideration it seems most probable that this condition on the radius of the star is well satisfied. Hence it would appear as if implosion of the central regions of such stars must imply explosion of the outer regions.

Two cases may be distinguished, leading to the occurrence of the r and p processes. A star with mass only slightly greater than Chandrasekhar's limit can evolve in the manner described above only after almost all nuclear fuels are exhausted. Hence any hydrogen present in the outer material must comprise at most only a small proportion of the total mass. This is the case of hydrogen deficiency that we associate with Type I supernovae, and with the operation of the r process. In much more massive stars, however, the central regions may be expected to exhaust all nuclear fuels and proceed to the point of implosion while much hydrogen still remains in the outer regions. Indeed the "central region" is to be defined in this connection as an innermost region containing a mass that exceeds Chandrasekhar's limit. For massive stars the central region need be only a moderate fraction of the total mass, so that it is possible for a considerable proportion of the original hydrogen to survive to the stage where central implosion takes place. This is the case of hydrogen excess that in the former paper we associated with the Type II supernovae, where the p process may occur. However, the rarity of the p-process isotopes, and hence the small amount of material which must be processed to synthesize them, suggests that if Type II supernovae are responsible, the p process is a comparatively rare occurrence even among them. On the other hand, in any supernova in which a large flux of neutrons was produced, a small fraction of those having very high energies might escape to the outer parts of the envelope, and after decaying there to protons, might interact with the envelope material and produce the p-process isotopes.

Energy from the explosive thermonuclear reactions, perhaps as much as fifty percent of it, will be carried inwards, causing a heating of the material of the central regions. It is to this heating that we attribute the emission of the elements of the iron peak during the explosion of supernovae.

C. Supernova Light Curves

The left-hand side of Plate 4 shows the Crab Nebula, the remnant of a supernova which exploded in 1054 A.D. The right-hand side of Plate 4 shows three exposures of the galaxy IC 4182 during the outburst and fading of the supernova of 1937. The first exposure (20 min) shows the supernova at maximum brightness; the second (45 min) shows it about 400 days after maximum; on the third (85 min) the supernova is too faint to be detected.

Throughout this paper we have proposed that the production of the e-process, r-process, and p-process isotopes takes place immediately preceding or during such outbursts of supernovae. This is the only type of astrophysical process that we know of, apart from an initial state in an evolutionary cosmology, in which it can be presumed that the necessary high-energy, rapidly evolving conditions can be expected. Thus while we were searching for such an astrophysical situation the discovery that the half-life for decay by spontaneous fission of Cf^{254} was equal to the half-life for decay of the light curve of the supernova in IC 4182 (Bu56, Ba56) led us to make the detailed calculations which are described in Secs. VII and VIII.

Since this original work was carried out, there has been further investigation of the half-life of Cf^{254} (see Sec. VIII). Also we have re-examined the evidence concerning the light curves of supernovae. Results of the calculations have allowed us to re-examine the basic assumption made in our earlier papers, that the energy released in the decay of Cf^{254} must dominate over all of the other possible sources of decay energy.

The half-life of 55 ± 1 days was obtained from Baade's result on the supernova in IC 4128. In Fig. XII, 1 we show Baade's light curve of the supernova in IC 4182, and the light curves of Tycho Brahe's nova and Kepler's nova (Ba43, Ba45), together with estimated points for the supernova of 1054 (the Crab Nebula) (Ma42). For SN IC 4182 the observations have been made for as long as 600 days after maximum. The absolute magnitude of this supernova at maximum has been

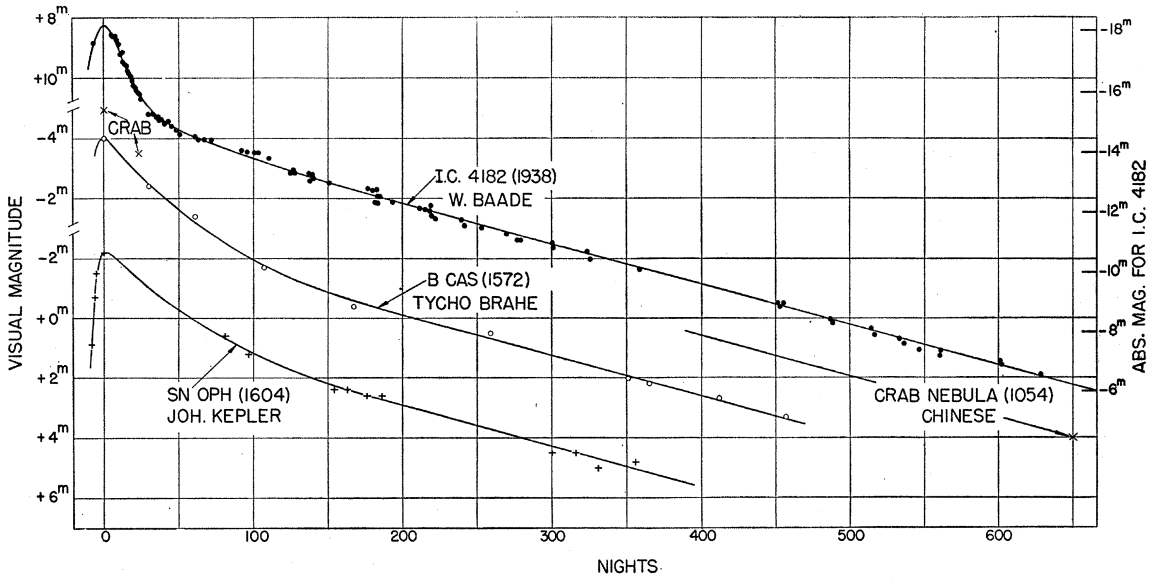

FIG. XII,1. Light curves of supernovae by Baade (Ba43, Ba45, Ba56). Measures for SN IC 4182 are by Baade; those for B Cassiopeiae (1572) and SN Ophiuchi (1604) have been converted by him to the modern magnitude scale from the measures by Tycho Brahe and Kepler. The three points for the supernova of 1054 are uncertain, being taken from the ancient Chinese records (Ma42). The abscissa gives the number of nights after maximum; the left-hand ordinate gives the apparent magnitude (separate scale for each curve); the points for the Crab Nebula belong on the middle scale, i.e., that for B Cassiopeiae. The right-hand ordinate gives the absolute magnitude for SN IC 4182 (Ba57a) derived by using the current distance scale. Compare with Fig. VIII,3.

estimated, with the best current value for the distance scale correction, to be -18.1 (Ba57a). The absolute magnitude of the Crab at its maximum in 1054 has been estimated to be -17.5 to -16.5 (To1355, Ma42, Ba57a). The value of 55 days for the half-life of SN IC 4182 was derived mainly from the points far out in the light curve where apparent magnitudes fainter than $+19$ had to be measured. Any systematic errors in measuring the faint magnitudes in this region would have the effect of changing the estimated value of this half-life and the possibility that such errors are present cannot be ruled out (Ba57a). The uncertainties in the measurements of other supernovae, which are not described here, are such that either all supernovae have true half-lives of 55 days, or they may have a unique half-life slightly different from 55 days, but lying in the range 45–65 days. Alternatively there may be an intrinsic variation between different light curves. These observational uncertainties must be borne in mind when considering the following discussion.

The only comparisons that we shall attempt to make will be between the total radiant energy emitted by a supernova and the form of the light curve, and the total energy emitted and the energy decay curves based on the calculations in Sec. VIII. No direct comparison between the early parts of the supernova light curve and the energy decay curve are possible, since the energy-degradation processes and the energy-transfer processes in the shell of the supernova will distort the relation between the two. Some aspects of this part of the problem have recently been studied by Meyerott and Olds (Me57).

Astrophysical arguments concerning the amounts of the r-process elements which are built in supernovae come from three different directions:

(i) The supernova or the supernovae which synthesized the r-process material of the solar system.

(ii) The Crab Nebula, the only remnant of a supernova which has been studied in any detail.

(iii) The light curves of supernovae which enable us to estimate how much material has been involved in the outburst, either on the assumption that the energy released by one or two of the isotopes dominates, and defines the light curve, or on the assumption that all of the decay activity is important.

Let us suppose that a total mass $mM_\odot$ is ejected in a particular supernova outburst and that a fraction f of this mass is in the form of r-process elements. Further, let us suppose that a fraction g of the heavy elements is in the form of Cf^{254}, or alternatively that a fraction g' of the heavy elements is in the form of Fe^{59}. In Sec. VIII it was shown that under specific conditions this was the only other isotope which could be expected to contribute an amount of energy comparable with Cf^{254}. Since it has a half-life of 45 days it cannot be *a priori* excluded from the discussion, particularly if, as has been stated above, some of the supernova light curves may have half-lives which are near to 45 days.

On the basis that Cf^{254} is responsible for the total

energy released in the light curve after this curve has reached its exponential form, an approximation which will be corrected in what follows, we have, since the energy release per Cf^{254} nucleus is 220 Mev, that the total energy available in the Cf^{254} is $gfm\times 1.67\times 10^{51}$ ergs. Of the unknowns m, f and g, m is unknown for all supernovae except the Crab which may have a mass of about 0.1 $M_\odot$ (Os57) so that $m\simeq 0.1$. The uncertainties in this value are discussed later. The value of f is determined by the conditions in the outer envelope of the star which have developed in its evolution to the supernova stage, while g is determined by the degree of building in the r process.

We consider the possible values of f. The most favorable situation that can be attained is reached when there are equal amounts by number of hydrogen, helium, and carbon, oxygen, and neon taken as a group, and about 1% of iron by number. In schematic terms, the sequence of events is as follows. The protons are captured by the C^{12}, O^{16}, Ne^{20} to form, after beta decay, C^{13}, O^{17}, Ne^{21}. Each of these now captures an alpha particle and releases a single neutron, so that we end with O^{16}, Ne^{20}, and Mg^{24}, together with a source of free neutrons. To build into the heavy element region, about 100 neutrons per Fe^{56} nucleus are required. These neutrons have come, via the steps that we have outlined, from the original hydrogen. Thus by number the necessary ratio of hydrogen to iron is 100:1, and the mass ratios of the material are hydrogen: helium: carbon: oxygen: neon: iron = 1:4:4:5.33:6.67:0.56, so that the iron comprises about 2% of the envelope by mass, and after the outburst about 6% of the mass is transformed into heavy elements. The maximum value of f is then 0.06. This situation is probably unrealistic because we have assumed essentially that the efficiency of the process is 1, since all of the hydrogen has been converted to neutrons, and each light nucleus has produced one neutron. As a more conservative value we will take $f_{\max}=0.01$.

A number of values of g are possible; $g=1$ corresponds to the situation in which all of the heavy elements are transmuted to Cf^{254}. We can see no reason from the standpoint of nuclear physics why this should occur and consequently we are inclined to disregard it.

If all of the r-process material is transformed into material lying in the range of atomic weight $230<A<260$, then we see from Fig. VII,4 that $g\simeq 0.04$. This is also a very unreal situation, since while no material resides in the region below $A\simeq 230$, it is also supposed that no fission has taken place to terminate the r process and hence to begin to cycle the material between $A=110$ and $A=260$.

We can consider a third situation in which, starting from iron, the material has been driven so that all of the r-process peaks have been produced in the way that has been described in Sec. VII. Further nuclear activity results in the following situation. The material begins to cycle between $A=110$ and $A=260$, because, following fission of the heaviest isotopes, the fission fragments help to build in the regions $A=110$ and $A=150$, and the total amount of material which resides in the region $110<A<260$ is determined by the amount of cycling which can take place. Sufficient cycling reduces the amount of material in the region <110, and in the extreme case we suppose that no iron is left and also no material is left in the $N=50$ peak. Under these conditions we estimate, from Figs. VII,3 and VII,4, that $g\simeq 0.011$. Here we have supposed that only half of the abundance follows the part of the cycle between $A\simeq 110$ and $A\simeq 150$.

A fourth situation is just that which has been supposed to give rise to the solar-system abundances of the r-process isotopes. In this case a steady state is reached and no appreciable cycling has taken place. Reference to Fig. VII,3 and VII,4 shows that $g\simeq 0.002$.

A fifth situation is that in which the r-process isotopes are built from a base material which is very much lighter than iron. In this case we have only been able to make a guess concerning the amount of material residing in r-process elements with $A\lesssim 45$. We then find that $g\simeq 0.0004$.

Thus the largest value of g which appears to be plausible is obtained when cycling takes place and $g=0.011$, and it is this value which we use in the following calculations.

There is a spread in absolute magnitudes at maximum of supernovae and probably also a spread in the difference between the maximum absolute magnitude and the magnitude at which the exponential form of the light curve sets in. Let us suppose that this difference is δM. Then the total energy emitted under the exponential tail of the light curve is given by

$$E_t=3.8\times 10^{33}\times 10^{0.4(4.6-M-\delta M)}\times 8.64\times 10^4\times \tau_m \text{ ergs},$$

where M is the absolute magnitude of the supernova at maximum and τ_m days is the mean life of the radioactive element responsible. The energy emitted by the sun and its absolute magnitude have been taken as 3.8×10^{33} erg/sec and 4.6, respectively. Thus for Cf^{254},

$$E_t=2\times 10^{42}\times 10^{0.4(-M-\delta M)}$$

where the half-life of $Cf^{254}=61$ days has been used. If the half-life is 56 days, then the value of E_t must be reduced by about 8%. We have shown in Sec. VIII that the total energy release parallels the spontaneous fission release over the time when the Cf^{254} predominates. Thus when making this calculation we must include the contribution which the other radioactive elements apart from Cf^{254} make to the curve. From Fig VIII,1 we find that the ratio of the total energy release to the Cf^{254} energy release is 8:7. On including this correction factor, we have that the total energy release from the mass $mM_\odot=gfm\times 1.91\times 10^{51}$ ergs. Thus putting $g=0.011$, $f=0.01$, we have that

$$m\times 1.07\times 10^{+5}=10^{0.4(-M-\delta M)}$$

or

$$\log m = 0.4(-M-\delta M)-5.029.$$

For the Crab, in which $M \simeq -17.5$ and δM is arbitrarily taken as $+3$,

$$m = 5.9,$$

while for SN IC 4182, in which $M=-18.1$ and $\delta M = +2.5$,

$$m = 16.1.$$

The mass of the Crab Nebula is estimated to be $0.1\ M_\odot$ (Os57). This result was based on the assumption that, although the hydrogen/helium ratio was abnormally small, in that their number densities were approximately equal, the remainder of the elements had normal abundances, as given, for example, in the appendix. The situation regarding the elements other than hydrogen and helium in the Crab is still very unclear. For example, there appear to be no identifications or information on iron, though we should expect that the iron/hydrogen or iron/helium ratio would be abnormally large. The mass is actually derived by obtaining the density of free electrons per unit of mean atomic weight, and finally by making the assumption about the composition described above. Thus the total mass estimated is approximately a linear function of the mean atomic weight which is assumed. Hence if the material in the shell had the initial composition which we consider to be typical of a highly-evolved star, the mean atomic weight would be about ten times that assumed in Osterbrock's calculation. Hence a mass $\simeq 1\ M_\odot$ would not be unreasonable. This leaves a discrepancy between the calculated and observed mass of this supernova shell of a factor of ~ 5. This may be partly or completely accounted for by uncertainties in estimating the absolute magnitude of the Crab supernova at maximum. Baade (Ba57a) has estimated that the absolute magnitude was -17.5 at maximum by using the observation that the star vanished from naked eye visibility about 650 days after maximum, and assuming that the total drop in brightness was the same as that for the supernova in IC 4182. Mayall and Oort (Ma42) have estimated that it reached -16.5 by using the observation that the supernova ceased to be visible in daylight after 23 days. If the supernova did reach only -16.5 at maximum, the mass needed to be ejected in the envelope is $\sim 2.4\ M_\odot$. An alternative explanation is that the Crab never did show that exponential light curve which is typical of supernovae in which Cf^{254} is produced.

For IC 4182, a mass of $\sim 10\ M_\odot$ is not unreasonable, particularly if it is supposed that the amount of mass ejected is roughly proportional to the maximum luminosity of the supernova.

It is of some interest to carry out a calculation similar to that made for Cf^{254} on the assumption that Fe^{59} is alone responsible for the light curve, particularly since we have earlier pointed out that this isotope will dominate the energy decay among the isotopes apart from Cf^{254}, and also since some supernova light curves have half-lives possibly nearer to 45 than to 55 days.

In this case the total energy released by Fe^{59} is given by

$$E_t = g' f m \times 8.6 \times 10^{49} \text{ ergs}.$$

The most favorable situation which can be envisaged is that as before there is about 2% by mass of iron in the envelope, so that $f \simeq 0.01$, but that there is a paucity of neutrons so that perhaps only 5 neutrons per iron nucleus are made available. On this assumption the proportion of the material which captures 3 neutrons to form Fe^{59} can easily be calculated, as indicated in Sec. VIII, if it is assumed that the capture cross section remains constant, and we find that about 20% of the original Fe^{56} will be transformed to Fe^{59}. In this case $g'=0.2$ (see Sec. VIII). Thus we obtain the result that

$$g' f m \times 4.44 \times 10^{49} = 1.48 \times 10^{42} \times 10^{0.4(M-\delta M)}$$

or

$$\log m = 0.4(-M-\delta M)-4.78.$$

This equation for m has a numerical constant which differs by a factor of 0.25 in the logarithm from that for the Cf^{254} decay. Thus we conclude that if the physical conditions described above are plausible, the large-scale production of Fe^{59} will be able to explain a light curve with a half-life of 45 days if the mass ejected is about 1.8 times greater than that demanded if Cf^{254} is responsible. However, if sufficient neutrons were available to build the r-process elements in the same proportions that they appear in the solar system the relative amount of Fe^{59} can only be estimated rather uncertainly, as has been done in Sec. VIII, but the value of g becomes much smaller than 0.2 and the mass required in the envelope becomes unreasonably large.

The relatively low efficiency of conversion of the material into either Cf^{254} or Fe^{59} arises from the small values of g, g', and f which seem to be plausible. The values of g and g' are determined by the nuclear physics of the situation. However, as previously stated, f is in essence determined by the initial composition of the envelope material. The maximum efficiency would be achieved when $f=1$, i.e., when the whole of the material was initially composed of iron and was changed to heavy elements in the supernova explosion. Such a condition does not seem very plausible. However, if a situation could arise in which some of the outer core material has been converted to isotopes in the iron peak by the equilibrium process (Sec. IV) while the inner core material has evolved to its last equilibrium configuration consisting of a pure neutron core, an instability in which the neutrons were mixed into the iron would result in the production of the r-process isotopes, and since there would be a plentiful supply of neutrons, cycling in the range $110 \lesssim A \lesssim 260$ would

take place. Under these conditions the total amount of mass demanded to produce the curves from Cf^{254} would be only about 0.06 $M_\odot$ in the case of the Crab, and only about 0.16 $M_\odot$ in the case of SN IC 4182.

Finally, mention must be made of the possible distortion of the light curve by the decay of other isotopes with longer half-lives than those responsible for the regions considered. The most important case appears to be that of Cf^{252}, which has a half-life of 2.2 years. From Fig. VIII,3 this should begin to make the curve flatten out after about 500 days. This flattening is very gradual and could well be masked by the small systematic errors in magnitude measurement which may be present in the light curve of the supernova in IC 4182. If no distortions occur through the energy transfer processes in the envelope at this late stage, very accurate measurements by photoelectric techniques of the very faint end of the next supernova light curve which can be followed may provide a good experimental test of this supernova theory. In the same way further observations may be able to determine whether all supernovae have exactly the same decay curves, a question which may also have some bearing on the problem of whether or not the *r*-process component of the solar-system material was built in a single supernova outburst.

D. Origin of the *r*-Process Isotopes in the Solar System

As mentioned in Sec. VII, the forms of the back sides of the abundance peaks of the *r*-process isotopes might suggest that the conditions obtaining in a single supernova were responsible for their synthesis. Though this conclusion remains highly uncertain at the present time, it is worthwhile considering its astrophysical implications.

A possible sequence of events in the origin of the *r*-process isotopes in the solar system might then be as follows. The isotopes which are made by the other processes have already been synthesized and a cloud of such material is present. Inside this cloud a supernova outburst takes place. About 10^{-2} $M_\odot$ is converted by this process to isotopes built in the *r* process, and this material is gradually diluted by the other material present. The material will be diluted first by expansion of the supernova shell into the surrounding medium and second by effects of galactic rotation and turbulence in the interstellar gas. Thus the total volume in which this occurs may be crudely estimated in the following way. The volume of a supernova shell (the Crab Nebula) is about 1 psc^3 after about 10^3 years. Since a shell probably decelerates, it may be supposed that after about 10^4 years its volume is about 25 psc^3 and its expansion velocity has reached a numerical value equal to the mean random velocity of the interstellar clouds.

One time-scale of interest is that for the formation of the solar system, i.e., 10^7-10^8 years, and one epoch of interest is a period of this length occurring some 5×10^9 years ago. However the more important time scale for this argument is the time in which the material can disperse, between the time at which the supernova outburst occurred and the epoch at which the solar system condensed. From the argument based on the U^{235}/U^{238} ratio this time interval is of the order of 10^9 years. At an epoch more than 5×10^9 years ago, it is probable that effects of galactic shear and of turbulent motion were different from their present values, so that no very good estimates of the total dilution volume can be made from this standpoint. On the other hand we can estimate the dilution factor by the following method.

On the basis of discussion in Sec. XII, we can suppose that the *r*-process isotopes originally comprised a total of about 2 percent of a solar mass in the supernova outburst. At present in the solar system the ratio by mass of the *r*-process isotopes to hydrogen is $\sim10^{-6}$ (Table II,1). Thus the amount of hydrogen into which the shell expanded was $\sim2\times10^4$ $M_\odot$. If we suppose that the mean density of this gas was about 10^{-24} g/cm^3, the total volume was $\sim4\times10^{61}$ cm^3 so that the dimension of this volume was about 75 psc. This is a fraction of approximately 10^{-6} of the total volume of the galaxy. The total number of supernovae that would explode in $\sim10^9$ years is $\sim3\times10^6$, so that we might expect that, on an average, 3 might have gone off in this volume, so that one only is entirely possible. It does not appear unreasonable from this point of view, therefore, that a single supernova has been responsible for all of the material built by the *r* process currently present in the solar system.

XIII. CONCLUSION

In Secs. III to X we have discussed the details of the various synthesizing processes which are demanded to produce the atomic abundances of all of the isotopes known in nature, while in Secs. II, XI, and XII we have attempted to describe the astrophysical theory and observations which are relevant to this theory.

It is impossible in a short space to summarize the advantages and disadvantages of a theory with as many facets as this. However, it may be reasonable to conclude as follows. The basic reason why a theory of stellar origin appears to offer a promising method of synthesizing the elements is that the changing structure of stars during their evolution offers a succession of conditions under which many different types of nuclear processes can occur. Thus the internal temperature can range from a few million degrees, at which the *pp* chain first operates, to temperatures between 10^9 and 10^{10} degrees when supernova explosions occur. The central density can also range over factors of about a million. Also the time-scales range between billions of years, which are the normal lifetimes of

stars of solar mass or less on the main sequence, and times of the order of days, minutes, and seconds, which are characteristic of the rise to explosion. On the other hand, the theory of primeval synthesis demands that all the varying conditions occur in the first few minutes, and it appears highly improbable that it can reproduce the abundances of those isotopes which are built on a long time-scale in a stellar synthesis theory.

From the standpoint of the nuclear physics of the problem, a major advance in the last few years has been the gradual realization through the interplay between experiment and theory that the helium-burning reaction $3He^4 \rightarrow C^{12}$ will take place in stellar interiors; theoretical work on stellar evolution has shown that in the interiors of red giant stars the conditions are right for such a reaction to take place. Another major advance has been the realization that under suitable stellar conditions the (α,n) reactions on certain nuclei can provide a source of neutrons; these are needed to synthesize the heavy isotopes beyond the iron peak in the abundance distribution and also to build some of the isotopes lighter than this.

The recent analysis of the atomic abundances (Su56) has enabled us to separate the isotopes in a reasonable scheme depending on which mode of synthesis is demanded. In particular, the identification of the *r*-process peaks was followed by the separation of the heavy isotopes beyond iron into the *s*-, *r*-, and *p*-process isotopes, and has enabled us to bring some order into the chaos of details of the abundance curve in this region. The identification of Cf^{254} in the Bikini test and then in the supernova in IC 4182 first suggested that here was the seat of the *r*-process production. Whether this finally turns out to be correct will depend both on further work on the Cf^{254} fission half-life and on further studies of supernova light curves, but that a stellar explosion of some sort is the seat of *r*-process production there seems to be little doubt.

From the observational standpoint, the gradual establishment in the last few years that there are real differences in chemical composition between stars is the strongest argument in favor of a stellar synthesis theory. Details of this argument, and the attempts to show that some stars are going through, or have gone through, particular synthesizing processes, while others are simply condensed out of material which has been processed earlier, and have not yet had time to modify any of their own material, have been given in Sec. XI.

Many problems remain. Although we have given a tentative account of the events leading up to a supernova outburst, we have no detailed dynamical theory of such an explosion. Neither can we explain the principal features of supernova spectra. Further, we do not yet understand the way in which stars evolve to this stage. In fact, the whole problem of stellar evolution beyond the red-giant stage is beset on the theoretical side by problems which are very difficult to handle with the present computational techniques. From the observational side we can estimate the time taken for a star to move through a particular evolutionary stage by making counts of the number of stars in that part of the HR diagram of a star cluster, but these diagrams alone cannot tell us, for example, whether the evolutionary track is in the sense of increasing or decreasing surface temperature at a particular point, i.e., from right to left or from left to right in the diagram.

The whole question of the chemical compositions of stars is complicated by the problem of mixing, since in many cases it cannot be supposed that the atmospheric composition of a star is identical with that of its interior. Also, attempts to determine compositions of very distant or very faint stars are a challenging problem. However, the term "universal abundance" will remain meaningless in the astronomical sense until a reasonable sample of the material in different parts of our own galaxy has been investigated, with the determination of abundances of the heavier elements. It might be pointed out here that on the basis of this theory there may be nothing universal in the isotope ratios. In many cases different isotopes of an element are built by different processes, and the isotope ratios which are observationally known are almost entirely obtained from solar-system material. Thus, for example, another solar system might have condensed out of material consisting mainly of hydrogen and gas ejected from stars which had gone through hydrogen burning, helium burning, the *s* process, the α process, and the *e* process, but not the *r* process. In this case, among the heavy elements the *s*-process isotopes would be present but the *r*-process isotopes would be absent, so that those elements which are predominantly built by the *r* process would be in very low abundance. Thus in such a solar system, the inhabitants would have a very different sense of values, since they would have almost no gold and, for their sins, no uranium.

The question of the composition of the material in other galaxies is not likely to be settled in the near future, except for coarse estimates. Perhaps the only sample of such material which we are able to obtain is contained in the extremely high-energy cosmic-ray particles ($\geq 10^{18}$ ev) which may have been accelerated and escaped from extragalactic nebulae (Ro57, Bu57b).

We have attempted to show that the interchange of material between stars and the interstellar gas and dust will be sufficient to explain the observed abundances relative to hydrogen. This attempt has been reasonably successful, though the aging problems have demanded that we do not consider that synthesis has gone on at a uniform rate since our galaxy condensed. It appears that successive generations of stars are demanded, if all eight processes are to contribute, although it is conceivable that, in a particular star,

one part or another can go through each process, ending with a catastrophic explosion.

We have made some attempt to explain possible modes of production of deuterium lithium, beryllium, and boron, but at present we must conclude that these are little more than qualitative suggestions.

In the laboratory, it is probable that the next steps to be taken are studies of the further reactions in helium burning, i.e., $C^{12}(\alpha,\gamma)O^{16}$ and $O^{16}(\alpha,\gamma)Ne^{20}$, work on the neutron-producing reactions, and particularly on neutron capture cross sections, and accurate work on nuclear masses and binding energies, particularly of the nuclei in the rare-earth region.

In astrophysical observations it is important that evidence for the operation of the *s* process and of the *r* process be sought spectroscopically. In giant stars where the *s* process may operate, the elements most overabundant should be those in general with a stable, "magic number" isotope having a closed shell of neutrons. These elements are strontium, yttrium, zirconium, barium, lathanum, cerium, praseodymium, neodymium, lead, and bismuth. All of these except bismuth have been found to be overabundant in the only such star analyzed up to the present time (Table XI,3). In supernova remnants, such as the Crab nebula, the *r* process may have operated to enhance the abundance of the following heavy elements: selenium, bromine, krypton, tellurium, iodine, xenon, caesium, osmium, iridium, platinum, gold, thorium, and uranium. The effect of the *r* process on the abundance of the iron-group elements and of the lighter elements from carbon to calcium is very difficult to estimate. The iron-group elements will be depleted by neutron capture in the original envelope material but material from the core (*e* process) injected during the explosion may yield an over-all enhancement in the final expanding shell. The over-all abundance of the lighter elements is not appreciably changed by the supernova explosion but the ratio to hydrogen and helium will rise to a value considerably greater than that in unevolved matter. This is because in our model we require that hydrogen and helium originally be comparable in abundance by number to carbon, oxygen, and neon, and then that they be depleted by the energy generation and the neutron-releasing processes.

An important key to the solution of problems in stellar nucleogenesis lies in the determination of relative isotopic abundances in stars. It is realized that this is a very difficult problem spectroscopically, but nonetheless every possible technique of isotope analysis should be brought to bear in this regard. Relative isotopic abundances are many times as informative as relative element abundances.

In concluding we call attention to the minimum age of the uranium isotopes derivable from their relative production in the *r* process. Our calculations place this minimum age at 6.6×10^9 years with a limit of error on the low side of at most 0.6×10^9 years. This small limit of error comes directly from the relatively short half-life of U^{235}. We feel that this minimum age is significantly greater than the currently accepted value for the geochemical age of the solar system, 4.5×10^9 years. It is, indeed, comparable with the ages assigned to the oldest globular clusters so far discovered. Since it is a minimum, it will not be unexpected if still older objects exist in our galaxy. In fact the stars discussed by Sandage and Eggen, mentioned in Sec. XII A, whose ages may be $\gtrsim 8\times10^9$ years, might be such examples.

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APPENDIX

We give in the appendix all of the information that we have been able to collect which is relevant to the synthesis problem. A detailed explanation of the format of this table is given in Sec. II C. The abundances (N) of the isotopes have been taken predominantly from the work of Suess and Urey (Su56), though in a few cases the solar abundances of Goldberg *et al.* (Go57) have been given in parentheses below those of Suess and Urey. The neutron capture cross sections (σ) have been discussed in Sec. V B, while the method of assignment of the nuclei among the eight synthesizing processes has been described in Sec. II B. The designation m means "magic," i.e., the nucleus has a closed shell of neutrons and thus a small neutron capture cross section in the *s* process. The designation m' means that the progenitor in the *r* process is "magic." When m is used as a superscript, as in Eu^{152m}, it designates an isomeric state. A process which returns material back down the neutron capture chain is designated by the word *returns*.

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
1	$_1H$	0.3 (0.1%th)	4.00×10^{10}		Primeval					
2	$_1H_1$	$\lesssim 2\times10^{-4}$ (0.1%th)	5.7×10^{6}		x					
3	$_1H_2(\beta^-)$	12.26 yr			τ_β too long					
3	$_2He_1$	5000 (n,p) (extr.)			H burning					
4	$_2He_2$		3.08×10^{9}		H burning					
5	$_2He_3(n)$	2×10^{-21} sec			returns					
6	$_2He_4(\beta^-)$	0.82 sec								
6	$_3Li_3$	2000 (n,α) (exp)	7.4		x, returns					
7	$_3Li_4$	0.33 (1%th)	92.6		x					
8	$_3Li_5(\beta^-)$	0.84 sec								
8	$_4Be_4(\alpha)$	$\sim2\times10^{-16}$ sec			returns					
9	$_4Be_5$	0.085 (1%th)	20		x	9	$_3Li_6(\beta^-)\rightarrow n, 2He^4$	0.17 sec		returns
10	$_4Be_6(\beta^-)$	2.7×10^{6} yr				10	$_5B_5$	1000 (n,α) (0.1%th)	4.5	returns
11	$_4Be_7(\beta^-)$	Not known								
11	$_5B_6$	<0.5 (1%th)	19.5		x					
12	$_5B_7(\beta^-)$	0.025 sec								
12	$_6C_6$	0.047 (1%th)	3.50×10^{6}		He burning					
13	$_6C_7$	0.009 (1%th)	3.92×10^{4}		H burning					
14	$_6C_8(\beta^-)$	5600 yr								
14	$_7N_7$	1.0 (n,γ) (1%th) 1.7 (n,p) (0.1%th)	6.58×10^{6}		H burning					
15	$_7N_8$	0.00024 (1%th)	2.41×10^{4}		H burning	15	$_6C_9(\beta^-)$	2.4 sec		
16	$_7N_9(\beta^-)$	7.4 sec								
16	$_8O_8$	0.01 (1%th)	2.13×10^{7}		He burning					
17	$_8O_9$	1.0 (n,α) (0.1%th)	8.00×10^{3}		H burning returns	17	$_7N_{10}(\beta^-)\rightarrow n, O^{16}$	4.14 sec		returns
18	$_8O_{10}$	0.0021 (1%th)	4.36×10^{4}		H burning					
19	$_8O_{11}(\beta^-)$	29 sec								
19	$_9F_{10}$	0.094 (1%th)	1600		H burning					
20	$_9F_{11}(\beta^-)$	11 sec								
20	$_{10}Ne_{10}$	0.02	7.74×10^{6}		He burning					
21	$_{10}Ne_{11}$	0.6 (also n,α)	2.58×10^{4}		H burning returns					
22	$_{10}Ne_{12}$	0.36 (1%th)	8.36×10^{5}	3.0×10^{5}	H burning					
23	$_{10}Ne_{13}(\beta^-)$	40 sec								
23	$_{11}Na_{12}$	3.5	4.38×10^{4}	1.5×10^{5}	H burning, s					
24	$_{11}Na_{13}(\beta^-)$	15.0 hr								
24	$_{12}Mg_{12}$	1.8	7.21×10^{5}	1.3×10^{6}	He burning, α					
25	$_{12}Mg_{13}$	6 (also n,α)	9.17×10^{4}	5.5×10^{5}	s					
26	$_{12}Mg_{14}$	2.0	1.00×10^{5}	2.0×10^{5}	$s(m)$					
27	$_{12}Mg_{15}(\beta^-)$	9.5 min								
27	$_{13}Al_{14}$	5.3 (exp)	9.48×10^{4}	5.0×10^{5}	$s(m)$					
28	$_{13}Al_{15}(\beta^-)$	2.30 min								
28	$_{14}Si_{14}$	1.8 ($\frac{1}{3}\times Al^{27}$)	9.22×10^{5}	1.7×10^{6}	α, $s(m)$					
29	$_{14}Si_{15}$	12 (also n,α)	4.70×10^{4}	5.6×10^{5}	s					
30	$_{14}Si_{16}$	8	3.12×10^{4}	2.5×10^{5}	s					
31	$_{14}Si_{17}(\beta^-)$	2.62 hr								
31	$_{15}P_{16}$	32	1.00×10^{4}	3.2×10^{5}	s					
32	$_{15}P_{17}(\beta^-)$	14.5 day								
32	$_{16}S_{16}$	12 (also n,α)	3.56×10^{5}	4.3×10^{6}	α, s					
33	$_{16}S_{17}$	36 (also n,α and n,p)	2.77×10^{3}	1.0×10^{5}	s					
34	$_{16}S_{18}$	12	1.57×10^{4}	1.9×10^{5}	s					
35	$_{16}S_{19}(\beta^-)$	87 days								
35	$_{17}Cl_{18}$	33	6670	2.2×10^{5}	s	36	$_{16}S_{20}$	12	51	$r(m)$
36	$_{17}Cl_{19}(\beta^-)$	3.1×10^{5} yr 100				36	$_{17}Cl_{19}(\beta^-)$	3.1×10^{5} yr		α, s
						36	$_{18}A_{18}$	15	1.26×10^{5} [1.9×10^{6}]	α, s s decay
						37	$_{18}A_{19}$(EC)	35 day		
37	$_{17}Cl_{20}$	10	2180	2.2×10^{4}	$s(m)$	37	$_{17}Cl_{20}$	10	2180 [2.2×10^{4}]	$s(m)$
38	$_{17}Cl_{21}(\beta^-)$	37.3 min								
38	$_{18}A_{20}$	2	2.4×10^{4}	4.8×10^{4}	$s(m)$					

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
39	$_{18}A_{21}(\beta^-)$	260 yr				39	$_{18}A_{21}(\beta^-)$	260 yr		
39	$_{19}K_{20}$	8	2940	2.4×10^4	$s(m)$	40	$_{18}A_{22}$	15	(+K^{40} decay)	s s decay
40	$_{19}K_{21}$ (EC, 11%, β^- 89%)	1.3×10^9 yr 100	0.38		s only	41	$_{18}A_{23}(\beta^-)$	1.82 hr		
41	$_{19}K_{22}$	25	219	5.5×10^3	s	41	$_{19}K_{22}$	25	219 [5.5×10^3]	s
42	$_{19}K_{23}(\beta^-)$	12.5 hr				40	$_{20}Ca_{20}$	4	4.75×10^4 [1.9×10^5]	$\alpha(m)$ s decay
42	$_{20}Ca_{22}$	12	314	3.8×10^3	s					
43	$_{20}Ca_{23}$	20	64	1.3×10^3	s					
44	$_{20}Ca_{24}$	8	1040	8.3×10^3	α, s					
45	$_{20}Ca_{25}(\beta^-)$	160 days								
45	$_{21}Sc_{24}$	16	2.8 (63)	45 (1.0×10^3)	s					
46	$_{21}Sc_{25}(\beta^-)$	85 days								
46	$_{22}Ti_{24}$	10	194 (290)	1.9×10^3 (2.9×10^3)	s only e	46	$_{20}Ca_{26}$	4	1.6	r only
47	$_{22}Ti_{25}$	12	189 (280)	2.3×10^3 (3.4×10^3)	s e, r					
48	$_{22}Ti_{26}$	6	1790 (2600)	10.7×10^3 (1.6×10^4)	α, s e, r	48	$_{20}Ca_{28}$	4	87.7	r only (m)
49	$_{22}Ti_{27}$	8	134 (200)	1.1×10^3 (1.6×10^3)	s e, r					
50	$_{22}Ti_{28}$	4	130 (190)	5.2×10^2 (7.6×10^2)	$s(m)$ r, e	50	$_{23}V_{27}$	10	0.55 (1.1)	e
51	$_{22}Ti_{29}(\beta^-)$	5.80 min				50	$_{24}Cr_{26}$	8	344 (1800)	e
51	$_{23}V_{28}$	30	220 (430)		$e(m)$					
52	$_{23}V_{29}(\beta^-)$	3.77 min								
52	$_{24}Cr_{28}$	8	6510 (3.3×10^4)		$e(m)$					
53	$_{24}Cr_{29}$	31	744 (3800)		e					
54	$_{24}Cr_{30}$	8	204 (1000)		e	54	$_{26}Fe_{28}$	29	3.54×10^4 (1.3×10^4)	$e(m)$
55	$_{24}Cr_{31}(\beta^-)$	3.6 min								
55	$_{25}Mn_{30}$	124	6850 (7900)		e					
56	$_{25}Mn_{31}(\beta^-)$	2.58 hr								
56	$_{26}Fe_{30}$	29	5.49×10^5 (2.1×10^5)		e					
57	$_{26}Fe_{31}$	96	1.35×10^4 (5100)		e					
58	$_{26}Fe_{32}$	29	1980 (760)		e	58	$_{28}Ni_{30}$	40	1.86×10^4 (1.7×10^4)	e
59	$_{26}Fe_{33}(\beta^-)$	45 day								
59	$_{27}Co_{32}$	248	1800 (2200)		e					
60	$_{27}Co_{33}(\beta^-)$	5.2 yr								
60	$_{28}Ni_{32}$	40	7170 (6600)		e					
61	$_{28}Ni_{33}$	104	342 (310)		e					
62	$_{28}Ni_{34}$	40	1000 (910)		e					
63	$_{28}Ni_{35}(\beta^-)$	80 yr				63	$_{28}Ni_{35}(\beta^-)$	80 yr		
63	$_{29}Cu_{34}$	48	146 (1500)	7.0×10^3 (7.2×10^4)	s	64	$_{28}Ni_{36}$	40	318 [1.3×10^4] (290) [(1.2×10^4)]	s
64	$_{29}Cu_{35}$ (EC 42%; β^+ 19%)	12.8 hr				64	$_{29}Cu_{35}$ (β^- 39%)	12.8 hr		
64	$_{28}Ni_{36}$	40(0.61)	318 (290)	7.8×10^3 (7.1×10^3)	s	64	$_{30}Zn_{34}$	49(0.39)	238 [4.5×10^3]	s only
65	$_{28}Ni_{37}(\beta^-)$	2.56 hr				65	$_{30}Zn_{35}$ (EC, β^+)	245 day		
65	$_{29}Cu_{36}$	48	66 (680)	3.2×10^3 (3.3×10^4)	s	65	$_{29}Cu_{36}$	48	66 [3.2×10^3] (680) [(3.3×10^4)]	s

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
66	$_{29}Cu_{37}(\beta^-)$	5.1 min								
66	$_{30}Zn_{36}$	49	134	6.6×10^3	s					
67	$_{30}Zn_{37}$	64	20.0	1.3×10^3	s					
68	$_{30}Zn_{38}$	49	90.9	4.5×10^3	s					
69	$_{30}Zn_{39}(\beta^-)$	52 min								
69	$_{31}Ga_{38}$	84	6.86	576	s					
70	$_{31}Ga_{39}(\beta^-)$	21 min								
70	$_{32}Ge_{38}$	82	10.4	849	s only	70	$_{30}Zn_{40}$	49	3.35	r only
71	$_{32}Ge_{39}$(EC)	12 day								
71	$_{31}Ga_{40}$	84	4.54	381	s					
72	$_{31}Ga_{41}(\beta^-)$	14.1 hr								
72	$_{32}Ge_{40}$	82	13.8	1130	s					
73	$_{32}Ge_{41}$	148	3.84	568	s					
74	$_{32}Ge_{42}$	82	18.65	1520	s	74	$_{34}Se_{40}$	96	0.649	p
75	$_{32}Ge_{43}(\beta^-)$	82 min								
75	$_{33}As_{42}$	240	4.0($\frac{2}{3}$)	640	sr					
76	$_{33}As_{43}(\beta^-)$	26.7 hr								
76	$_{34}Se_{42}$	96	6.16	591	s only	76	$_{32}Ge_{44}$	82	3.87	r only
77	$_{34}Se_{43}$	360	5.07		rs					
78	$_{34}Se_{44}$	96	16.0		$r(m')$	78	$_{36}Kr_{42}$	107	0.175	p
79	$_{34}Se_{45}(\beta^-)$	7×10^4 yr 360				79	$_{34}Se_{45}(\beta^-)$	7×10^4 yr		
80	$_{34}Se_{46}$	96	33.8		$r(m')$	79	$_{35}Br_{44}$	480	6.78	$r(m')$ s decay
81	$_{34}Se_{47}(\beta^-)$	18 min				80	$_{35}Br_{45}$ (β^- 92%; EC 5%; β^+ 3%)	18 min		
81	$_{35}Br_{46}$	480	6.62		$r(m')$	80	$_{36}Kr_{44}$	107	1.14 [122]	s only
82	$_{35}Br_{47}(\beta^-)$	36 hr				81	$_{36}Kr_{45}$(EC)	2×10^5 yr 440		
82	$_{36}Kr_{46}$	107	5.90	630	s only	82	$_{36}Kr_{46}$	107	5.90 [630]	s only
83	$_{36}Kr_{47}$	440	5.89		$r(m')$	82	$_{34}Se_{48}$	96	5.98	r only (m')
84	$_{36}Kr_{48}$	107	29.3		$r(m')$	84	$_{38}Sr_{46}$	48	0.106	p
85	$_{36}Kr_{49}(\beta^-)$	10.4 yr								
85	$_{37}Rb_{45}$	112	4.73		$rs(m')$					
86	$_{37}Rb_{49}(\beta^-)$	18.6 day								
86	$_{38}Sr_{48}$	48	1.86	89	s only	86	$_{36}Kr_{50}$	2.4	8.94	r only (m)
87	$_{38}Sr_{49}$	60	1.33	80	s only	87	$_{37}Rb_{50}(\beta^-)$	4.3×10^{10} yr 9	1.77	r only (m)
88	$_{38}Sr_{50}$	4.8	15.6	75	$s(m)$					
89	$_{38}Sr_{51}(\beta^-)$	54 day								
89	$_{39}Y_{50}$	19	8.9	170	$s(m)$					
90	$_{39}Y_{51}(\beta^-)$	64.0 hr								
90	$_{40}Zr_{50}$	12	28.0	336	$s(m)$					
91	$_{40}Zr_{51}$	24	6.12	147	s					
92	$_{40}Zr_{52}$	18	9.32	118	s	92	$_{42}Mo_{50}$	24	0.364	$p(m)$
93	$_{40}Zr_{53}(\beta^-)$	9×10^5 yr 24				93	$_{40}Zr_{53}(\beta^-)$	9×10^5 yr		
94	$_{40}Zr_{54}$	18	9.48	121	s	93	$_{41}Nb_{52}$	160	1.00	r, s decay
95	$_{40}Zr_{55}(\beta^-)$	65 day				94	$_{41}Nb_{53}{}^m(\beta^-)$	6.6 min		
						94	$_{41}Nb_{53}(\beta^-)$	2×10^4 yr		
95	$_{41}Nb_{54}(\beta^-)$	35 day				94	$_{42}Mo_{52}$	240	0.226	p, s
95	$_{42}Mo_{53}$	324	0.382	124	s	95	$_{42}Mo_{53}$	324	0.382 [124]	s
96	$_{42}Mo_{54}$	240	0.401	96	s only	96	$_{40}Zr_{56}$	6.0	1.53	r only
						96	$_{44}Ru_{52}$	432	0.0846	p
97	$_{42}Mo_{55}$	324	0.232	75	s					
98	$_{42}Mo_{56}$	240	0.581 ($\frac{1}{2}$)	70	$s\approx r$	98	$_{44}Ru_{54}$	432	0.0331	p
99	$_{42}Mo_{57}(\beta^-)$	67 hr								
99	$_{43}Tc_{56}(\beta^-)$	2.1×10^5 yr 120				99	$_{43}Tc_{56}(\beta^-)$	2.1×10^5 yr		
100	$_{43}Tc_{57}(\beta^-)$	16 sec				99	$_{44}Ru_{55}$	1160	0.191	sr s decay
100	$_{44}Ru_{56}$	432	0.189	82	s only	100	$_{44}Ru_{56}$	432	0.189 [82]	s only
101	$_{44}Ru_{57}$	1160	0.253		rs	100	$_{42}Mo_{58}$	240	0.234	r only

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
102	$_{44}Ru_{58}$	432	0.467($\frac{1}{2}$)	100	$r\approx s$	102	$_{46}Pd_{56}$	504	0.0054	p
103	$_{44}Ru_{59}(\beta^-)$	40 day								
103	$_{45}Rh_{58}$	760	0.214		r					
104	$_{45}Rh_{59}(\beta^-)$	42 sec								
104	$_{46}Pd_{58}$	504	0.0628	32	s only	104	$_{44}Ru_{60}$	432	0.272	r only
105	$_{46}Pd_{59}$	880	0.1536		r					
106	$_{46}Pd_{60}$	504	0.1839		r	106	$_{48}Cd_{58}$	432	0.0109	p
107	$_{46}Pd_{61}(\beta^-)$	7×10^6 yr				107	$_{46}Pd_{61}(\beta^-)$	7×10^6 yr		
		880				107	$_{47}Ag_{60}$	1000	0.134	r
										s decay
108	$_{46}Pd_{62}$	504	0.180		r	108	$_{47}Ag_{61}$ (β^-, 98.5%; EC, 1.5%)	2.3 min		
109	$_{46}Pd_{63}(\beta^-)$	13.6 hr				108	$_{48}Cd_{60}$	432	0.0079	p, s
						109	$_{48}Cd_{61}$(EC)	1.3 yr		
109	$_{47}Ag_{62}$	1000	0.126		r	109	$_{47}Ag_{62}$	1000	0.126	r
110	$_{47}Ag_{63}(\beta^-)$	24 sec								
110	$_{48}Cd_{62}$	432	0.111	48	s only	110	$_{46}Pd_{64}$	504	0.0911	r only
111	$_{48}Cd_{63}$	1160	0.114($\frac{1}{2}$)	66	$r\approx s$					
112	$_{48}Cd_{64}$	432	0.212($\frac{1}{2}$)	46	$r\approx s$	112	$_{50}Sn_{62}$	84	0.0134	p
113	$_{48}Cd_{65}$	1160	0.110($\frac{1}{2}$)	64	$r\approx s$	113	$_{48}Cd_{65}{}^m(\beta^-)$	5 yr		
						113	$_{49}In_{64}$	1320	0.0046	ps
114	$_{48}Cd_{66}$	432	0.256($\frac{1}{2}$)	55	$r\approx s$	114	$_{49}In_{65}(\beta^-)$	72 sec		
115	$_{48}Cd_{67}(\beta^-)$	54 hr				114	$_{50}Sn_{64}$	84	0.0090	ps
115	$_{49}In_{66}(\beta^-)$	1320 (6×10^{14} yr)	0.105($\frac{1}{2}$)	70	$s\approx r$					
115	$_{49}In_{66}{}^m(\beta^-)$	4.5 hr				115	$_{50}Sn_{65}$	280	0.00465	ps
116	$_{49}In_{67}(\beta^-)$	13 sec								
116	$_{50}Sn_{66}$	84	0.189	16	s only	116	$_{50}Sn_{66}$	84	0.189 [16]	s only
117	$_{50}Sn_{67}$	280	0.102	29	s	116	$_{48}Cd_{68}$	432	0.068	r only
118	$_{50}Sn_{68}$	84	0.316	27	s					
119	$_{50}Sn_{69}$	280	0.115	32	s					
120	$_{50}Sn_{70}$	84	0.433	36	s	120	$_{52}Te_{68}$	156	0.00420	p
121	$_{50}Sn_{71}(\beta^-)$	~28 hr								
121	$_{51}Sb_{70}$	348	0.141	49	s					
122	$_{51}Sb_{71}(\beta^-)$	2.8 day								
122	$_{52}Te_{70}$	156	0.115	18	s only	122	$_{50}Sn_{72}$	84	0.063	r only
123	$_{52}Te_{71}$	520	0.0416	22	s only	123	$_{51}Sb_{72}$	348	0.105	r only (m')
124	$_{52}Te_{72}$	156	0.221	34	s only	124	$_{50}Sn_{74}$	84	0.079	r only (m')
						124	$_{54}Xe_{70}$	360	0.00380	p
125	$_{52}Te_{73}$	520	0.328		$r(m')$					
126	$_{52}Te_{74}$	156	0.874		$r(m')$	126	$_{54}Xe_{72}$	360	0.00352	p
127	$_{52}Te_{75}(\beta^-)$	9.3 hr								
127	$_{53}I_{74}$	880	0.80		$r(m')$					
128	$_{53}I_{75}$(β^- 95%)	25.0 min				128	$_{53}I_{75}$(EC+β^+ 5%)	25.0 min		
128	$_{54}Xe_{74}$	360	0.0764	28	s only	128	$_{52}Te_{76}$	156	1.48	$r(m')$
						129	$_{52}Te_{77}(\beta^-)$	72 min		
129	$_{54}Xe_{75}$	1400	1.050		$r(m')$	129	$_{53}I_{76}(\beta^-)$	1.7×10^7 yr		
						130	$_{53}I_{77}(\beta^-)$	12.6 hr		
130	$_{54}Xe_{76}$	360	0.162	58	s only	130	$_{54}Xe_{76}$	360	0.162 [58]	s only
131	$_{54}Xe_{77}$	1400	0.850		r	130	$_{52}Te_{78}$	156	1.60	r only (m')
						130	$_{56}Ba_{74}$	37.2	0.00370	p
132	$_{54}Xe_{78}$	360	1.078		r	132	$_{56}Ba_{76}$	37.2	0.00356	p
133	$_{54}Xe_{79}(\beta^-)$	5.27 day								
133	$_{55}Cs_{78}$	480	0.456		rs					
134	$_{55}Cs_{79}(\beta^-)$	2.3 yr								
134	$_{56}Ba_{78}$	37.2	0.0886	3.3	s only	134	$_{54}Xe_{80}$	360	0.420	r only
135	$_{56}Ba_{79}$	124	0.241($\frac{2}{3}$)	20	sr					
136	$_{56}Ba_{80}$	37.2	0.286	11	s only	136	$_{54}Xe_{82}$	4.9	0.358	r only (m)
						136	$_{58}Ce_{78}$	82	0.0044	p
137	$_{56}Ba_{81}$	124	0.414($\frac{2}{3}$)	34	sr					
138	$_{56}Ba_{82}$	10	2.622	26	$s(m)$	138	$_{57}La_{81}$(EC, β^-)	57.6 (2×10^{11} yr)	0.0018	p
						138	$_{58}Ce_{80}$	82	0.00566	p
139	$_{56}Ba_{83}(\beta^-)$	85 min								
139	$_{57}La_{82}$	32	2.00	66	$s(m)$					
140	$_{57}La_{83}(\beta^-)$	40.2 hr								
140	$_{58}Ce_{82}$	19	2.00	38	$s(m)$					

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
141	$_{58}Ce_{83}(\beta^-)$	32 day								
141	$_{59}Pr_{82}$	60	0.40	24	$s(m)$					
142	$_{59}Pr_{83}(\beta^-)$	19.1 hr								
142	$_{60}Nd_{82}$	36	0.39	15	s only (m)	142	$_{58}Ce_{84}$	82	0.250	r only
143	$_{60}Nd_{83}$	180	0.175	33	s					
144	$_{60}Nd_{84}$	120	0.344	42	s	144	$_{62}Sm_{82}$	72	0.0108	$p(m)$
145	$_{60}Nd_{85}$	180	0.119	21	s					
146	$_{60}Nd_{86}$	120	0.248	30	s					
147	$_{60}Nd_{87}(\beta^-)$	11.6 day								
147	$_{61}Pm_{86}(\beta^-)$	2.6 yr								
147	$_{62}Sm_{85}$	4800	0.100		rs					
148	$_{62}Sm_{86}$	1440	0.0748	108	s only	148	$_{60}Nd_{88}$	120	0.0824	r only
149	$_{62}Sm_{87}$	4800	0.0920		rs					
150	$_{62}Sm_{88}$	1440	0.0492	71	s only	150	$_{60}Nd_{90}$	120	0.0806	r only
151	$_{62}Sm_{89}(\beta^-)$	80 yr				151	$_{62}Sm_{89}(\beta^-)$	80 yr		
		4800								
152	$_{62}Sm_{90}$	1440	0.176		rs	151	$_{63}Eu_{88}$	5600	0.0892	rs
						152	$_{63}Eu_{89}$(EC, 72%; β^-, 28%)	13 yr or 9.2 hr		
153	$_{62}Sm_{91}(\beta^-)$	47 hr				152	$_{63}Eu_{89}{}^{m}(\beta^-$, 78%; EC, 22%)	13 yr or 9.2 hr		
						152	$_{64}Gd_{88}$	2400	0.00137	ps
						153	$_{64}Gd_{89}$(EC)	236 day		
153	$_{63}Eu_{90}$	5600	0.0976		rs	153	$_{63}Eu_{90}$	5600	0.0976	rs
154	$_{63}Eu_{91}(\beta^-)$	16 yr								
154	$_{64}Gd_{90}$	2400	0.0147	35	s only	154	$_{62}Sm_{92}$	1440	0.150	r only
155	$_{64}Gd_{91}$	6800	0.101		r					
156	$_{64}Gd_{92}$	2400	0.141		r	156	$_{66}Dy_{90}$	1320	0.00029	p
157	$_{64}Gd_{93}$	6800	0.107		r					
158	$_{64}Gd_{94}$	2400	0.169		r	158	$_{66}Dy_{92}$	1320	0.000502	p
159	$_{64}Gd_{95}(\beta^-)$	18 hr								
159	$_{65}Tb_{94}$	6000	0.0956		r					
160	$_{65}Tb_{95}(\beta^-)$	72 day								
160	$_{66}Dy_{94}$	1320	0.0127	17	s only	160	$_{64}Gd_{96}$	2400	0.149	r only
161	$_{66}Dy_{95}$	4400	0.105		r					
162	$_{66}Dy_{96}$	1320	0.142		r	162	$_{68}Er_{94}$	1116	0.000316	p
163	$_{66}Dy_{97}$	4400	0.139		r					
164	$_{66}Dy_{98}$	1320	0.157		r	164	$_{68}Er_{96}$	1116	0.00474	p? Too abundant
165	$_{66}Dy_{99}(\beta^-)$	2.32 hr								
165	$_{67}Ho_{98}$	2800	0.118		r					
166	$_{67}Ho_{99}(\beta^-)$	27.2 hr								
		>30 yr								
166	$_{68}Er_{98}$	1116	0.104		r					
167	$_{68}Er_{99}$	1840	0.0770		r					
168	$_{68}Er_{100}$	1116	0.0850		r	168	$_{70}Yb_{98}$	1080	0.00030	p
169	$_{68}Er_{101}(\beta^-)$	9.4 day								
169	$_{69}Tm_{100}$	1400	0.0318		r					
170	$_{69}Tm_{101}(\beta^-)$	129 day								
170	$_{70}Yb_{100}$	1080	0.00666	7.2	s	170	$_{68}Er_{102}$	1116	0.0470	r only
171	$_{70}Yb_{101}$	1200	0.0316($\frac{1}{3}$)	19	$s \approx r$					
172	$_{70}Yb_{102}$	1080	0.0480($\frac{2}{3}$)	35	sr					
173	$_{70}Yb_{103}$	1200	0.0356($\frac{2}{3}$)	29	sr					
174	$_{70}Yb_{104}$	1080	0.0678($\frac{2}{3}$)	49	sr	174	$_{72}Hf_{102}$	520	0.00078	p
175	$_{70}Yb_{105}(\beta^-)$	4.2 day								
175	$_{71}Lu_{104}$	1040	0.0488($\frac{2}{3}$)	34	sr	176	$_{70}Yb_{106}$	1080	0.0278	r only
176	$_{71}Lu_{105}{}^{m}(\beta^-)$	3.7 hr				176	$_{71}Lu_{105}(\beta^-)$	Large σ (7.5×10^{10} yr)	0.0013	p
176	$_{72}Hf_{104}$	520	0.0226	12	s only	177	$_{71}Lu_{106}(\beta^-)$	6.8 day		
177	$_{72}Hf_{105}$	1040	0.0806($\frac{2}{3}$)	56	sr	177	$_{72}Hf_{105}$	1040	0.0806	sr
178	$_{72}Hf_{106}$	520	0.119	62	s					
179	$_{72}Hf_{107}$	1040	0.0604	63	s					
180	$_{72}Hf_{108}$	520	0.155	81	s	180	$_{74}W_{106}$	480	0.0006	p
181	$_{72}Hf_{109}(\beta^-)$	46 day								
181	$_{73}Ta_{108}$	1320	0.065	86	s					
182	$_{73}Ta_{109}(\beta^-)$	112 day								
182	$_{74}W_{108}$	480	0.13	62	s					
183	$_{74}W_{109}$	1640	0.070($\frac{2}{3}$)	77	sr					
184	$_{74}W_{110}$	480	0.15	72	s	184	$_{76}Os_{108}$	1176	0.00018	p
185	$_{74}W_{111}(\beta^-)$	74 day								
185	$_{75}Re_{110}$	2040	0.0500($\frac{1}{3}$)	51	$s \approx r(m')$					

A	Main line	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	σN	Process	A	Weak line or bypassed	$\sigma(n,\gamma)$ in mb or $\tau_{\frac{1}{2}}$	N	Process
186	$_{75}Re_{111}$ (β^- ~95%)	91 hr				186	$_{75}Re_{111}$ (EC ~5%)	91 hr		
186	$_{76}Os_{110}$	1176	0.0159	19	*s* only	186	$_{74}W_{112}$	480	0.14	*r*(*m'*)
						187	$_{74}W_{113}(\beta^-)$	24 hr		
187	$_{76}Os_{111}$	2480	0.0164	41	*s* only	187	$_{75}Re_{112}(\beta^-)$	2040 (~5×10^{10} yr)	0.0850	*r*(*m'*)
						188	$_{75}Re_{113}(\beta^-)$	17 hr		
188	$_{76}Os_{112}$	1176	0.133		*r*(*m'*)	188	$_{76}Os_{112}$	1176	0.133	
189	$_{76}Os_{113}$	2480	0.161		*r*(*m'*)					
190	$_{76}Os_{114}$	1176	0.264		*r*(*m'*)	190	$_{78}Pt_{112}$	384	0.0001	*p*
191	$_{76}Os_{115}(\beta^-)$	16 day								
191	$_{77}Ir_{114}$	3160	0.316		*r*(*m'*)					
192	$_{77}Ir_{115}$ (β^- 96.5%)	74 day				192	$_{77}Ir_{115}$ (EC 3.5%)	74 day		
192	$_{78}Pt_{114}$	384	0.0127	4.9	*s* only	192	$_{76}Os_{116}$	1176	0.410	*r*(*m'*)
193	$_{78}Pt_{115}$	long				193	$_{76}Os_{117}(\beta^-)$	31 hr		
193	$_{78}Pt_{115}{}^{(m)}$(EC)	3.4 day								
193	$_{77}Ir_{116}$	3160	0.505		*r*(*m'*)	193	$_{77}Ir_{116}$	3160	0.505	*r*(*m'*)
194	$_{77}Ir_{117}(\beta^-)$	19 hr								
194	$_{78}Pt_{116}$	384	0.533		*r*(*m'*)					
195	$_{78}Pt_{117}$	1240	0.548		*r*(*m'*)					
196	$_{78}Pt_{118}$	384	0.413		*r*	196	$_{80}Hg_{116}$	240	0.00045	*p*
197	$_{78}Pt_{119}(\beta^-)$	19 hr								
197	$_{79}Au_{118}$	1200	0.145		*r*					
198	$_{79}Au_{119}(\beta^-)$	2.70 day								
198	$_{80}Hg_{118}$	240	0.0285	6.8	*s* only	198	$_{78}Pt_{120}$	384	0.117	*r* only
199	$_{80}Hg_{119}$	680	0.0481		*rs*					
200	$_{80}Hg_{120}$	240	0.0656($\frac{2}{3}$)	11	*sr*					
201	$_{80}Hg_{121}$	680	0.0375		*rs*					
202	$_{80}Hg_{122}$	240	0.0844($\frac{2}{3}$)	14	*sr*					
203	$_{80}Hg_{123}(\beta^-)$	48 day								
203	$_{81}Tl_{122}$	276	0.0319($\frac{2}{3}$)	6	*sr*					
204	$_{81}Tl_{123}$ (β^- ~98%)	4.1 yr				204	$_{81}Tl_{123}$ (EC ~2%)	4.1 yr		
204	$_{82}Pb_{122}$	50	0.0063 (0.26)	0.32 (13)	*s* only	204	$_{80}Hg_{124}$	240	0.0194	*r*
205	$_{82}Pb_{123}$(EC)	~5×10^7 yr 100				205	$_{80}Hg_{125}(\beta^-)$	5.2 min		
						205	$_{81}Tl_{124}$	276	0.0761($\frac{2}{3}$) [14]	*sr* *s* decay
						206	$_{81}Tl_{125}(\beta^-)$	4.20 min		
206	$_{82}Pb_{124}$	25	0.122($\frac{1}{2}$) (5.0)	1.5 (62)	*s* cycles *r* decay	206	$_{82}Pb_{124}$	25	0.122($\frac{1}{2}$) (5.0) [1.5] [(62)]	*s* cycles *r* decay
207	$_{82}Pb_{125}$	50	0.0995($\frac{1}{2}$) (4.1)	2.5 (100)	*s* cycles *r* decay					
208	$_{82}Pb_{126}$	10	0.243 (10)	2.4 (100)	*s*(*m*) cycles (*r* decay)					
209	$_{82}Pb_{127}(\beta^-)$	3.3 hr								
209	$_{83}Bi_{126}$	15	0.144(?)	1.2(?)	*s*(*m*) cycle (*r* decay)					
210	$_{83}Bi_{127}{}^m$ (β^- 56%) (RaE)	5.0 day				210	$_{83}Bi_{127}$ (α 44%)	(2.6×10^6 yr) 5		
210	$_{84}Po_{126}(\alpha)$	138.40 day			(*m*)	211	$_{83}Bi_{128}(\alpha)$	2.15 min		
206	$_{82}Pb_{124}$	cycles				207	$_{81}Tl_{126}(\beta^-)$	4.78 min		
						207	$_{82}Pb_{125}$	cycles		

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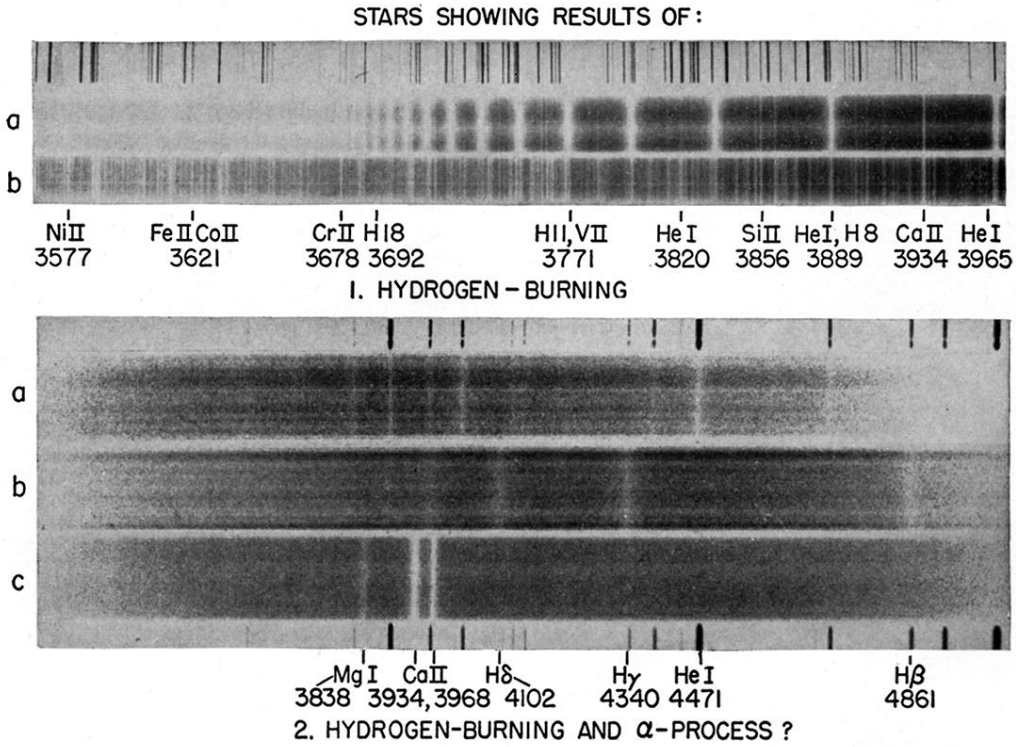

PLATE I.

PLATE 1. Portions of the spectra of stars showing the results of hydrogen burning and possibly the α process. Upper: (a) Normal A-type star, η Leonis, showing strong Balmer lines of hydrogen and a strong Balmer discontinuity at the series limit. (b) Peculiar star, υ Sagittarii, in which hydrogen has a much smaller abundance than normal. Lower: (a) White dwarf, L 1573–31, in which hydrogen is apparently absent. The comparison spectrum above the star is of a helium discharge tube; note the lines of helium in the star's spectrum. (b) White dwarf, L 770–3, which shows broad lines due to hydrogen only, for comparison with (a) and (c). (c) White dwarf, Ross 640, which shows only the two lines due to Ca II and a feature due to Mg I. All the spectrograms in this plate were obtained by J. L. Greenstein; the upper two are McDonald Observatory plates and the lower three are Palomar Observatory plates.

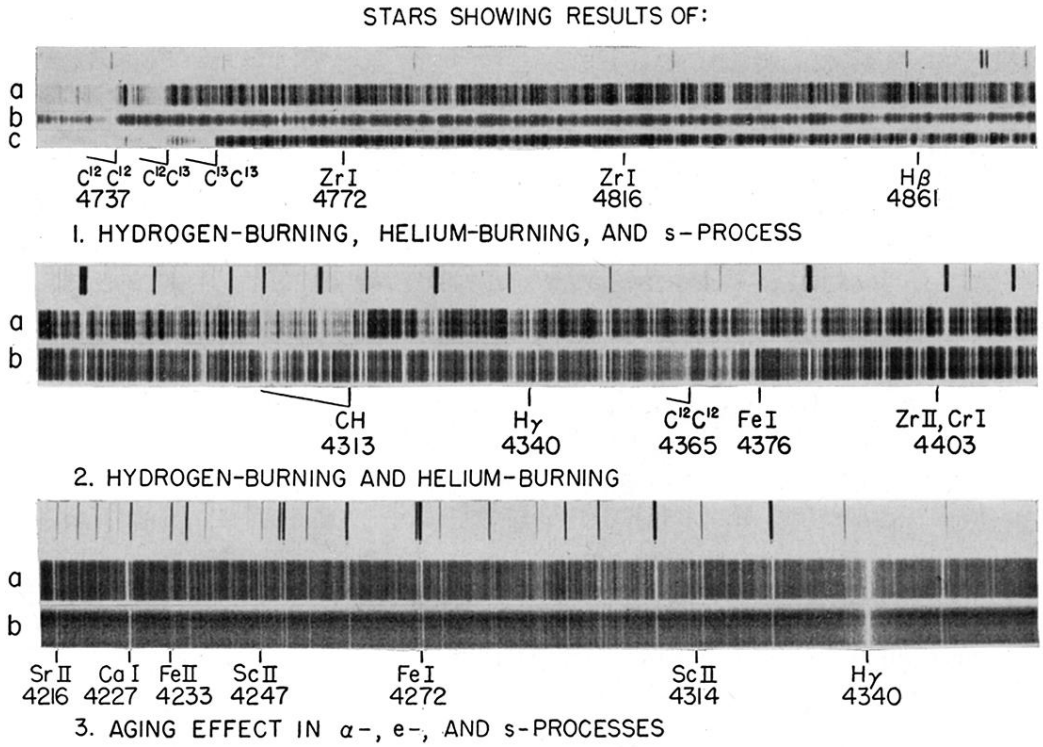

PLATE 2.

PLATE 2. Portions of the spectra of stars showing different aspects of element synthesis. Upper: (a) Normal carbon star, X Cancri, which has $C^{12}/C^{13}\sim 3$ or 4. (b) Peculiar carbon star, HD 137613, which shows no C^{13} bands, and in which hydrogen is apparently weak. (c) Normal carbon star, HD 52432, which has $C^{12}/C^{13}\sim 3$ or 4. Note that ZrI lines appear to be strongest in (a). Middle: (a) Normal carbon star, HD 156074, showing the CH band and Hγ. (b) Peculiar carbon star, HD 182040, in which CH is not seen, although the weak band of C_2 at λ 4365 is visible. Hγ is also very weak, indicating that hydrogen has a low abundance. Lower: (a) Normal *F*-type star, ξ Pegasi. (b) Peculiar star, HD 19445, which has a slightly lower temperature than ξ Pegasi, yet all lines but hydrogen are much weakened, showing that the abundances of α-, *e*-, and *s*-process elements are much lower than normal ("aging" effect). The middle two spectra were obtained by J. L. Greenstein, the remainder, with the exception of HD 137613, by E. M. and G. R. Burbidge.

STARS SHOWING RESULTS OF s-PROCESS

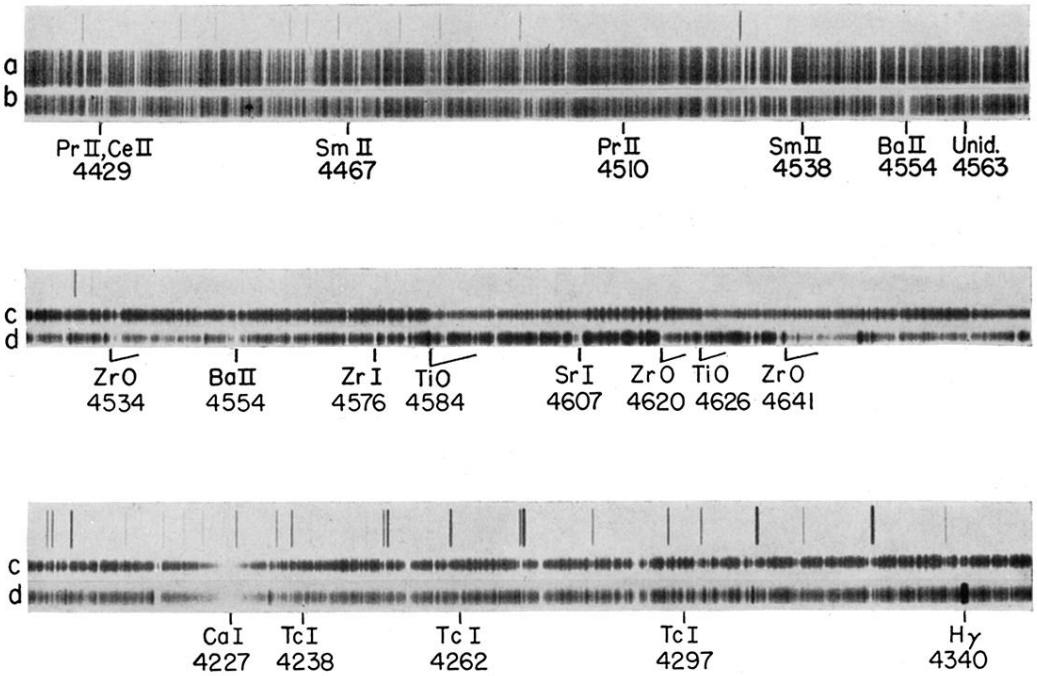

PLATE 3.

PLATE 3. Portions of the spectra of stars showing the results of the *s* process. Upper: (a) Normal *G*-type star, κ Geminorum. (b) Ba II star, HD 46407, showing the strengthening of the lines due to the *s*-process elements barium and some rare earths. Middle: (c) *M*-type star, 56 Leonis, showing TiO bands at λλ 4584 and 4626. (d) *S*-type star, R Andromedae, showing ZrO bands which replace the TiO bands. Lines due to Sr I, Zr I, and Ba II are all strengthened. Lower: (c) Another spectral region of the *M*-type star, 56 Leonis; note that Tc I lines are weak or absent. (d) R Andromedae; note the strong lines of Tc I. The spectrum of R Andromedae was obtained by P. W. Merrill, and the upper two spectra by E. M. and G. R. Burbidge.

Crab Nebula

Supernova in IC 4182

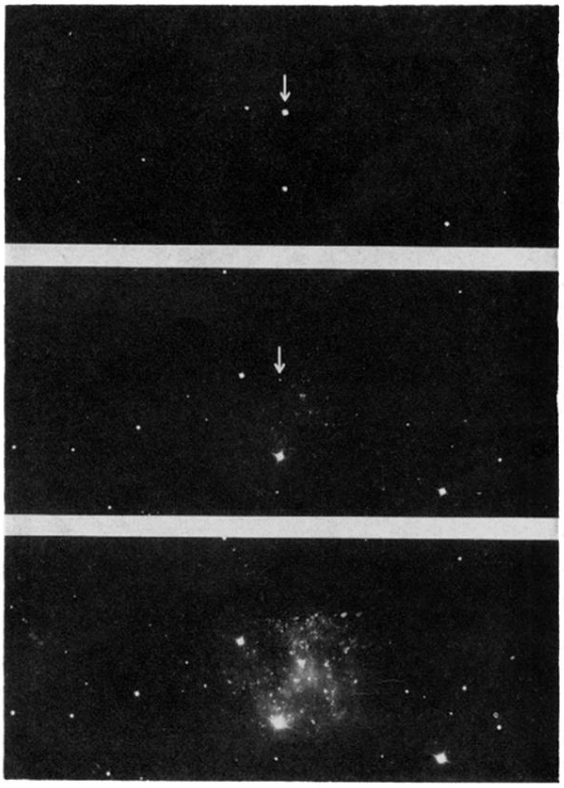

PLATE 4. Left: The Crab Nebula, photographed in the wavelength range λ6300–λ6750. The filamentary structure stands out clearly at this wavelength, which comprises light mainly due to the Hα line. Right: The supernova in IC 4182, photographed (a) September 10, 1937 at maximum brightness—exposure 20 m; (b) November 24, 1938, about 400 days after maximum—exposure 45 m; (c) January 19, 1942, about 1600 days after maximum, when the supernova was too faint to be detected—exposure 85 m. Note that the lengths of the three exposures are different. These plates were taken by W. Baade, to whom we are indebted for permission to reproduce them.